Proceedings for the 35th Annual Conference of the Society for Astronomical Sciences

SAS-2016

The Symposium on Telescope Science

Editors:
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Jerry L. Foote
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June 16-18, 2016
Ontario, CA
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Published by the Society for Astronomical Sciences, Inc.
Rancho Cucamonga, CA

First printing: June 2016

Photo Credits:

Front Cover:
NGC 2024 (Flame Nebula) and B33 (Horsehead Nebula)
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Welcome to the 35<sup>th</sup> annual Symposium of the Society for Astronomical Sciences! This year’s agenda reflects the broad diversity of interests among SAS participants, with papers covering photometry, spectroscopy, interferometry and astrometry; instruments ranging from eyeballs to CCDs and spectrographs to radios; and projects ranging from education to citizen-science to a variety of astronomical research targets.

It takes many people to have a successful conference, starting with the Program Committee. This year the regular committee members are:

Robert Gill  Robert D. Stephens
Cindy Foote  Jerry Foote
Robert Buchheim  Dale Mais
Wayne Green

We thank the staff and management of the Ontario Airport Hotel for their efforts to accommodate the Society and our activities.

Membership dues and Registration fees do not fully cover the costs of the Society and the annual Symposium. We owe a great debt of gratitude to our corporate sponsors: Sky and Telescope, Woodland Hills Camera and Telescopes, PlaneWave Instruments, Santa Barbara Instruments Group/Cyanogen, DC-3 Dreams; and Sierra Remote Observatories. Thank you!

Finally, there would be no Symposium without the speakers and poster presenters, the attentive audience, and the community of researchers and educators who apply their small telescopes to research activities. We thank all of you for making the SAS Symposium one of the premiere events for professional-amateur collaboration in astronomy.

Robert K. Buchheim
Jerry L. Foote
Dale Mais

2016 May
Symposium Sponsors

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The Role of Amateur Astronomers in Exoplanet Research

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Abstract

Because of recent technological advances in imaging equipment and processing software, research grade characterization of exoplanet transits is now within the realm of amateur astronomers. This paper will describe the current state of exoplanet observing by amateur astronomers and the best practices used for conducting such observations. In addition, a pipeline will be described that has proved successful in obtaining high quality, exoplanet science data from these observations. The paper will also describe a major collaboration currently underway between a world-wide network of amateur astronomers and a Hubble Space Telescope (HST) science team. Finally, the growing need for exoplanet observations by amateur astronomers will be discussed, as well as an example of their role in characterizing other “exo-objects.”

1. Introduction

With the recent availability of sensitive and low-noise, yet affordable, CCD detectors and sophisticated software processing, amateur astronomers are now able to conduct research-grade exoplanet science with amazing accuracy. This has therefore enabled amateur astronomers to work closely with professional astronomers to detect new exoplanets, to confirm candidate exoplanets, and to provide refined ephemeris for the detailed study of known exoplanets.

Although typically using the transit method to conduct such observations, a small number of amateur astronomers have also successfully used Doppler spectroscopy to conduct radial velocity measurements of planetary systems. In addition to exoplanet observing, amateur astronomers are now even assisting professional astronomers in the detection and characterization of large asteroids (planetesimals) orbiting their host stars.

Section 2 of this paper will provide a brief background on the history of exoplanet observing by amateur astronomers. Section 3 will discuss some underlying concepts that are fundamental to such observations. Section 4 will describe best practices of exoplanet observing by amateur astronomers. Section 5 will discuss a pipeline that has proven useful for conducting exoplanet transit modeling. Section 6 will review an on-going professional/amateur collaboration involving an HST study of the atmosphere of some 15 exoplanets. Section 7 will discuss the increasing need for amateur astronomer involvement in exoplanet research.

2. Background

In 1995, 51 Pegasi b was the first exoplanet detected around a main sequence star. Since then, over 3,200 exoplanets have been confirmed by Kepler and other space and ground-based observatories.

Amateur astronomers have been successfully detecting exoplanets for at least a decade, and have been doing so with amazing accuracy! Furthermore, they have been able to make such observations with the same equipment that they use to create fabulous looking deep sky pictures or variable star light curves.

Several examples exist of amateur astronomers providing valuable data in support of exoplanet research. In 2004, a team of professional/amateur astronomers collaborated on the XO Project, which resulted in the discovery of several exoplanets. The KELT (Kilodegree Extremely Little Telescope) program currently uses a world-wide network of amateur astronomers and small colleges, along with professional astronomers, to conduct follow-up observations of candidate exoplanets transiting bright stars (Pepper et al., 2007). As described later in this paper, a network of amateur astronomers is currently supporting a Hubble survey of some 15 exoplanets by conducting observations in the optical wavelength, while Hubble is studying these same exoplanets in the near-infrared. Amateur astronomer observations such as this help to refine the ephemeris of already known exoplanets. The mere fact that astronomers can accurately model the transits of existing exoplanets means that it is also theoretically possible for them to discover new exoplanets! For example, by detecting variations in the transit time of a known exoplanet (a technique called “transit time
variations”, or TTV), amateur astronomers can detect the existence of another planet orbiting the host star.

The formalization of techniques for exoplanet detection by amateur astronomers began in 2007 with Bruce Gary’s publication of “Exoplanet Observing for Amateurs” (Gary, 2014). At the same time, Gary began an attempt to archive the exoplanet observations of other amateur astronomers. This archive, the Amateur Exoplanet Archive (AXA), was subsequently transferred to the now more active Exoplanet Transit Database (ETD) project, an online archive sponsored by the Czech Astronomical Society (Poddany et al., 2010). An effort is currently underway by the American Association of Variable Star Observers (AAVSO) to support archiving of exoplanet observations in its AAVSO International Database (AID).

More recently, the author has developed “A Practical Guide to Exoplanet Observing” (Conti, 2016). It includes best practices for exoplanet observing and is intended for both the newcomer to exoplanet observing, as well as for the more-experienced exoplanet observer. In addition, it includes a step-by-step guide for using AstroImageJ (AIJ), freeware software that is an all-in-one package for image calibration, differential photometry, and exoplanet modeling (Collins et al., 2016).

3. Exoplanet Observing Concepts

Unlike deep sky imaging that is interested in an aesthetically pleasing picture, the dominant method used by amateur astronomers doing exoplanet observing involves taking precise measurements of changes in the brightness of a target star (the terms “target star” and “host star” are used interchangeably in this paper). This is done in order to detect the transit of a planet across the face of that star.

The “brightness” of a star – that is, its measured flux – is a function of many variables. For example, the apparent brightness of a star vs. its actual brightness is affected by the interstellar dust through which photons from the star must pass. This is true for both ground-based and space-based observatories. However, the effects on measured flux are much more severe for ground-based observatories due to several other variables. The dominant variables that affect ground-based flux measurements include the atmospheric mass through which a measurement is taken, as well as transient effects such as thin clouds and changing atmospheric turbulence. Even local affects at the observing location such as heat rising from surrounding land masses or building structures will affect a star’s measured flux. Technically, the measured flux of a star is the number of photons detected per area per unit of time. In practice, the cumulative ADU (Analog Digital Unit) counts of a group of CCD photosites is used as a proxy for flux.

3.1 Differential Photometry

Because of these variables affecting ground-based observations, a method is needed to differentiate the changes in the star’s flux due to these effects vs. a drop in the star’s light due to a transiting object such as an exoplanet. A method used to differentiate such changes in flux is differential photometry. As the name implies, differential photometry measures the difference in flux of the target star relative to one or more comparison (“comp”) stars. In theory, a change due to any of the aforementioned variables will equally affect both the target star and comp stars and therefore any change in the relative flux of the target star is usually due to a local occurrence on or surrounding that star. It is assumed, or course, that the comp stars are themselves not variable. Thus, selection of appropriate comp stars is an important part of the differential photometry process.

The basic concept of exoplanet observing involves taking a series of images of the field surrounding the host star of a suspected exoplanet before, during, and after the predicted times of the exoplanet transit across the face of its host star. The flux from the host star, as well as potential comp stars, are recorded for later analysis.

Exoplanet transits are typically two (2) hours or longer. However, it is desirable that there also be a sufficient amount of out-of-transit (OOT) time during the imaging session in order to establish a good baseline for later model building. The desired amount of OOT time is 1 hour before the beginning of the transit to 1 hour after. Thus, an exoplanet imaging session can be as long as 4 hours or more.

The measured flux of a target or comp star from a ground-based observatory is a combination of the inherent brightness of the star and a contribution from background sky glow. The effects of moon glow, light pollution, and reflected ground light contribute to this background sky component. In order to determine the measured flux of a star, this background sky contribution needs to be factored out. This is done through the use of an “annulus,” which captures the background sky glow around the star and an “aperture” which captures the flux of the star (see Figure 1). The contribution of the former then is “subtracted” from the latter in order to get a better representation of the star’s measured brightness.
This aperture/annulus combination is then applied to the target star and one or more comparison stars. Differential photometry software will then process each image taken during the imaging session and measure the flux of the target star and selected comp stars after compensating for the measured background sky around each star. The difference between the measured flux of the target star and the comp star(s) is then calculated for each image. In some cases, this “difference” is a subtraction of the fluxes, in other cases it is the measured flux of the target star divided by the combined flux of the comp stars. This relative change in flux of the target star then theoretically represents effects local to the target star itself.

3.2 Transit Modeling

The data points that represent the target star’s change in flux are then used to create a “best fit” transit model. This best fit results in estimates of key parameters about the exoplanet and its transit. These parameters include:

1. the square of the ratio of the radius of the exoplanet \((R_p)\) to that of its host star \((R_\ast)\),
2. the ratio of the exoplanet’s semi-major orbital radius \((a)\) to \(R_\ast\),
3. the center point \(T_c\) and the duration of the transit,
4. the inclination of the exoplanet’s orbit relative to the observer’s line-of-sight.

Thus, by knowing the radius \(R_\ast\) of the exoplanet’s host star, the exoplanet observer can actually then estimate the radius of the exoplanet, as well as the radius of its semi-major orbit. Seager et al. (2003) describes the derivation of important planetary system parameters from an exoplanet transit light curve.

4. Best Practices for Exoplanet Observing

With the increasing number of exoplanet observations now being conducted by amateur astronomers, a set of best practices can now be established. When applied to exoplanet transit observations, the following principles have resulted in a better fit of the data collected:

**Image scale:** The image scale (i.e., arc-seconds per pixel) of the imaging system, after any binning of the CCD camera is considered, should be such that the full width at half maximum (FWHM) of the host star spans three (3) or more pixels. Unlike deep sky imaging where the imager is interested in pinpoint stars, exoplanet observing is more interested in collecting accurate information about the flux of the exoplanet’s host star and one or more comparison stars. If because of the imaging system’s image scale it is not possible to achieve this desired pixel span, then it is acceptable for the observer to defocus the image such that this span of 3 or more pixels can be achieved. Defocusing may also help increase the chance of finding suitable comp stars. A point-spread-function (PSF) resulting from defocusing of 10-20 pixels is acceptable, unless the sky background is high. However, one should be aware that defocusing could cause a neighboring star to blend into a host or comp star aperture.

**Selection of comp stars:** The comp stars used should be as close in magnitude as possible to the exoplanet’s host star. Ideally the ensemble of comp stars should be a mix of ones that are 0.5-1.5 times the brightness (flux) of the host star, which translates to 0.75 greater in magnitude (i.e., dimmer) than the host star to 0.44 less in magnitude (i.e., brighter) than the host star. Also, because full exoplanet transits will typically take place over a range of elevations, the brightness of stars of different stellar type will increase (decrease) differently as AIRMASS decreases (increases). Therefore, it is also best to choose comp stars of similar stellar type. However, since some exoplanet modeling software, such as AJI, is able to “detrend” the effects of AIRMASS, choosing comp stars of similar brightness to the host star is more important than choosing stars of similar stellar type. Above all, comp star(s) should not be a variable star. A good source for such information is the AAVSO’s Variable Star Plotter utility (AAVSO, 2016).

**Flat fielding:** No matter what method is used to create flat field frames (electroluminescence panels, twilight
flats, dome flats, etc.), the result should be a uniform flat field. This is especially true in the case of German equatorial mounts where a meridian flip would cause the target and comp star(s) to land on different parts of the CCD detector.

**Autoguiding:** Because it is nearly impossible to achieve a flat field that perfectly corrects an imaging system’s pixel-to-pixel sensitivity differences, it is imperative that the exoplanet observer minimize the movement of the host and comp stars on the CCD detector. This is best achieved by making sure that the observer’s mount is properly polar aligned and has minimal periodic error. Most importantly, however, autoguiding is needed to make sure that the field stays as stationary as possible on the detector throughout the time of the observation.

**Filter use:** If the results are going to be part of a professional/amateur collaboration effort, a specific photometric filter will most likely be requested that the observer should use.

**Time synchronization:** The observer should have software running on his/her image capture computer that frequently synchronizes the computer’s clock to the U.S. Naval Observatory’s Internet time server. Dimension 4 is an example of such freeware that runs on Windows computers (see Dimension4, 2016). The update period for such clock synchronizations should be set to at least every 2 hours.

**Time system:** Because of the existence of several different time systems, the exoplanet observer should be aware which one is being used for the transit prediction, which one is being entered into the image FITS headers, which one is being used during the light curve modeling process, etc. The more commonly used time systems are:

- Julian Date/Universal Coordinated Time (JD UTC)
- Heliocentric Julian Date/Universal Coordinated Time (HJD UTC)
- Barycentric Julian Date/Barycentric Dynamical Time (BJD TDB).

If the results of the exoplanet observation are to be used in a professional/amateur collaboration, BJD TDB would be the desired time standard to use during the model fitting process.

5. Exoplanet Processing Pipeline

An ideal image processing and modeling pipeline, especially one used for science purposes, should be one that meets at least these requirements:

- It should be efficient in terms of its use of computing resources.
- It should be user friendly.
- It should be repeatable, that is, given the same raw image and calibration data, an independent party should be able to produce nearly identical results.

This section describes a pipeline for capturing, processing, and modeling exoplanet transits that is believed to meet these requirements. It has evolved and been refined after its application to dozens of exoplanet transit observations over several months.

The pipeline described below is separated into three phases:

- Preparation Phase
- Image Capture and Calibration Phase
- Photometry, Plotting, and Modeling Phase.

5.1 Preparation Phase

5.1.1 Information Collection

An Excel spreadsheet should be used that allows an exoplanet observer to record certain pieces of critical information during each Phase and provides a convenient source of information needed by the modeling software. If no changes are made to the observer’s instruments or location, many of the items can be used across multiple observing sessions. The latest version of such a worksheet can be downloaded from the author’s website (Conti, 2016).

5.1.2 Considerations in Selecting an Exoplanet Target

The following are useful sources for predicting exoplanet transits for a given time period at a particular observer’s location:

- NASA Exoplanet Archive (NASA, 2016)
- Exoplanet Transit Database (ETD, 2016)
- Extrasolar Planet Transit Finder (Coughlin, 2016).

If the exoplanet observer is selecting his/her own exoplanet target (i.e., one not specified as part of a particular research campaign), then the following selection criteria should be considered that will result in a more satisfying result:

- Beginning time of transit – since it is desirable that the imaging session start 1 hour before the beginning of the transit, this may negate some exoplanet candidates since this might put the start time during twilight.
- Duration of transit – with some transit durations longer than others, the observer may want to pick a candidate whose total session time (considering the desire to image 1 hour after the actual end of the transit) is suitable to the observer.
- Magnitude of the host star and depth of the transit – exoplanet host stars can range in V magnitude from 8.0 to over 13.0, and dips in the star’s magnitude due to the exoplanet transit can range from thousandths to hundredths of a magnitude. The observer might therefore want to choose an exoplanet target with a larger predicted % drop in magnitude (i.e., depth divided by star magnitude) than another potential target.

5.1.3. Meridian Flip Predictions

For observers with German equatorial mounts, the observer should predict approximately when, if at all, a meridian flip might be required during the imaging session. This prediction is typically done using the observer’s navigation software and is helpful so that the observer can be available during the meridian flip to make any necessary adjustments for repositioning the imaging system’s field-of-view as expeditiously as possible.

5.1.4. Choice of Exposure Time

Today’s CCD detectors typically have a linear range up to a point, after which they become saturated, and at which point any additional photons hitting the CCD photosite will not be registered. Thus, it is critical that the target and comp stars never reach saturation. It is important that the exposure time be chosen such that a decent SNR is achieved, but not long enough that saturation occurs. Furthermore, if the target star is predicted to rise toward the local meridian and, therefore its light will pass through less and less air mass, saturation could possibly be reached. Thus, the observer should also take this into consideration when choosing an exposure time. Note that defocusing could also help this situation.

In order to initially set the correct exposure time, a series of test images should be taken with increasing exposure time. The SNR of the target star, as well as its ADU counts, could then be measured for each exposure setting. An exposure setting that maximizes SNR, but doesn’t present a potential for saturation during the imaging session should then be considered as the ideal exposure time.

5.1.5. File Directories

On the computer that runs the observer’s image capture software, the following subdirectories should initially be setup: Bias, Darks, Flats, Test Images, and Science Images. Note: here the term Science Images refers to the raw images of the field-of-view containing the exoplanet host star; such images are also often referred to as Lights by some image processing software. Finally, a subdirectory named Analysis should be setup where measurement and model fit files from the analysis software can be stored. These subdirectory names will be used below; however, if other names are preferred for these subdirectories, then such names can be substituted for their respective counterparts.

5.1.6. Stabilization of Imaging System to Appropriate Temperature

The imaging system should be put in place with enough time for it to reach its desired temperature set-point, which might also require enabling of its cooling system.

5.1.7. Generation of Flat Files

Whether twilight flats are taken or flats are generated by using an electroluminescence panel, they should be redone (ideally) prior to or after each imaging session using the same imaging chain as was used for taking the Science Images, and with the imaging chain not having been displaced or moved.

5.1.8. Autoguiding

Autoguiding should be used during the Image Capture Phase below, unless the observer’s mount is of such accuracy that it can maintain guiding within a few pixels for the duration of the transit observation (where the actual value of “a few” depends on the FWHM of the host star). See item 4. of Section 4 for the reasons why autoguiding is so important. The observer’s autoguiding mechanism should be calibrated, if not yet done or if the autoguiding software does not automatically correct for changes in declination. If needed, calibration should also be done for any active optics (AO) system that is being used.

5.2 Image Capture and Calibration Phase

The observer’s normal image capture software is used to capture the Science Images into the Science Images subdirectory during this phase. Should a
meridian flip be necessary during the imaging session, then at the time the meridian flip is needed, the observer should:
1. Abort the image capture software
2. Stop autoguiding
3. Execute the meridian flip
4. Reposition the target star as necessary in the camera’s field-of-view (plate-solving software may be of help here)
5. Enable autoguiding
6. Enable the image capture software.

Prior to or after the capture of Science Images, the observer would also capture a series of darks, flat field, and bias calibration files. A rule of thumb is to capture an odd number (but no less than 17) of images for each such calibration series. An odd number is suggested, since this better allows for a median combine to later be used to create the master dark, master flat, and/or master bias files. Guidelines for exposure times for each of these calibration file types are as follows:
1. dark files – exposure time should equal that of the Science Images and should be taken at the same temperature as were the Science Images;
2. flat field files – exposure time is dependent on the flat fielding technique used, but typically takes 3 seconds or less; however, this exposure time may have to be increased for cameras with automatic shutters so that no shutter shading occurs; a general rule of thumb is that the resulting histogram of the flat field should be at the midpoint of the CCD’s dynamic range;
3. bias files – a bias file is a dark file of 0 second exposure time.

Typically, flat dark files are also created since the flat fields themselves contain dark current that needs to be subtracted out. When used, flat darks are taken at the same exposure time as the flat field files themselves. However, some calibration software can scale the Master Dark to the same exposure time as the Master Flat, thereby obviating the need for the observer to generate flat dark files.

5.3 Photometry, Plotting, and Modeling Phase

5.3.1. Photometry

After the Science Images have been calibrated, aperture photometry is then applied to the target star and several comp stars. The initial aperture/annulus radii sizes are determined according to the following guidelines:

1. The initial radius of the aperture (r₁) should be at least 2 times the number of FWHM pixels.
2. The initial radius of the inner annulus (r₂) should be chosen such that it and the radius of the outer annulus below create an annulus region that excludes any other stars that happen to be close to the target star. It should be noted that some aperture photometry software, such as AIJ, will automatically attempt to ignore the pixels of stars in the annulus region.
3. The initial value of the outer annulus radius should equal the SQRT(4*r₁²+r₂²). This should produce an annulus that contains 4 times the number of pixels as are in the aperture.

Note: It may be necessary to first align the images before the above aperture photometry step can take place.

Using the above aperture/annulus rings, a table of measurements is created across all of the calibrated Science Images. It may be necessary to delete some of the images from consideration because of elongated stars as a result of movement of the CCD detector due to wind or other uncorrectable guiding issues. Each entry in this measurement table generally consists of the date/time of the observation, the change in flux of the target star relative to the comp stars, and an error estimate of this measurement. Other entries may include estimated AIRMASS at the time of the observation, the change in flux of each comp star to the other comp stars, and their error estimates.

5.3.2. Plotting of Results

At this point, several data plots are useful in assessing the overall quality of the observation and the choice of comp stars. In the AIJ example depicted in Figure 2, the top curve represents the flux of the target star relative to the total counts of the selected comp stars. As easily seen, this shows that a transit occurred during the observational session. The second curve (in red) shows the transit model fit to this relative flux, with the third scatter plot showing the differences between the model fit and the observed relative flux.
The black, green and purple plots in Figure 2 depict each comp star’s change in flux relative to the other comp stars. These plots would show if there is any variability in the comp stars themselves. Some exoplanet transit modeling programs, such as AJ, allow the user to deselect comp stars from the ensemble if a comp star shows variability, or if a particular comp star is having a dominant effect on the relative flux calculation of the host star.

Finally, a plot of predicted AIRMASS (blue curve) vs. the total ADU counts of the comp stars (brown dots) is indicative of any passing clouds or other atmospheric conditions. In the example depicted in Figure 2, there was very good transparency since the total comp star ADU counts matched the changes in AIRMASS throughout the observing session.

5.3.3. Modeling

Depending upon the exoplanet transit modeling software, several “priors” are usually input by the user of the modeling software. These include one or more of the following:

1. period of the exoplanet’s orbit
2. radius of the host star
3. predicted ingress and egress times
4. limb darkening coefficients (either linear or quadratic).

Such information is available from a variety of sources. For most confirmed exoplanets, the website http://www.exoplanets.org provides information on the first three items above. Limb darkening coefficients, which are a function of characteristics of the host star and the type of filter used, can be found at Ohio State (2016).

If a meridian flip has occurred, or if there is a break in the image session and the stars have moved to a different part of the CCD detector, then some modeling software can adjust for resulting shifts in the computed relative flux values. In addition, other parameters such as AIRMASS, total comp star ADU counts, position on the detector of the host star can also be used to “detrend” the raw relative flux values before any model fit takes place.

Figure 3 is an example of an exoplanet fit based on data collected by amateur astronomer Paul Benni for exoplanet Wasp-76b. It shows how a Meridian Flip detrend parameter was used by AJ to take care of an offset in data points caused by the target and comp stars landing on a different part of the CCD detector after the meridian flip occurred.

Figure 4 shows sample results of an AJ exoplanet model fit. The Best Fit values are the resulting model outputs and the RMS value is a measure of the “goodness” of the model fit. Also shown is how AIRMASS was selected as one of the detrend parameters for this particular model fit.
Amateur astronomers are currently contributing to a major Hubble investigation of exoplanet atmospheres. This program is entitled “Metallicity and Cloud Survey of Exoplanetary Atmospheres Prior to JWST (James Webb Space Telescope)” (STScI, 2015). The program’s principal investigator is noted planetary scientist Dr. Drake Deming. The purpose of the survey is to observe some 15 exoplanets ranging from gas giants to planets of two Earth masses, and is intended to address two fundamental questions. They are, as stated in the survey team’s proposal:

1. “Using the water molecule as a probe, we will investigate the degree to which planetary envelopes are enriched in heavy elements as a function of planetary mass, and how that enrichment might be affected by mass loss.

2. We will define the degree to which clouds occur in exoplanetary atmospheres, over a wide range in temperature, surface gravity, and stellar irradiation.”

The Hubble team is using HST’s WFC3 (Wide-Field Camera 3) to obtain high precision spectroscopy of the subject planets. This is done using a “grating prism” (grism) and measurements taken in the 1.4 μm near-infrared band. Observations began November 26, 2015 and will continue throughout the rest of 2016.

The survey will make one or more HST “visits” to the survey’s 15 exoplanet targets. Each visit will consist of five (5) orbits of Hubble around the Earth. Because Hubble’s orbit is every 97 minutes, Hubble can’t continuously observe a full transit since its observations are occulted by the Earth a portion of the time.

Amateur astronomers are observing the same exoplanets as in the Hubble survey. The purpose of these ground-based observations is two-fold:

1. Refine the ephemeris of the subject planets, and
2. Determine any unusual stellar activity such as star spots or flares on the planet’s host star.

These ground-based observations by amateur astronomers are taking place in the optical spectrum. Standard photometric filters are preferred so that observations can be compared and since limb-darkening coefficients are available for most of the exoplanets for a given filter.

For each observation of an exoplanet that looks promising, the author does a further analysis of the data. Assuming that the results look favorable, a subset of the data is passed along to Joey Rodriguez, a member of the KELT team. When several favorable observations for a given exoplanet have been obtained, Rodriguez conducts a “global fit” that produces the refined ephemeris data. These results are then passed on to the Hubble co-investigator responsible for that particular exoplanet.

Currently over 50 observers from around the world are participating in this collaboration. In addition, members of the KELT follow-up network are also participating. Observers are needed across different time zones since transit opportunities occur at predicted times, which are only visible at night from certain regions. Of course even in these regions, local weather conditions and the elevation (i.e., air mass) through which a given target rises or falls during the observing session could affect the overall quality of the observation.

Participants in this Hubble collaboration range from backyard amateur astronomers with modest-size (11” and larger) aperture telescopes to those with access to remote, dark sites. The equipment and software requirements are almost identical to that needed for typical deep sky imaging.

In order to standardize the methodology by which the participants contribute to the program, a Standard Methodology document was developed. In addition, as Hubble is within a month or so of its next exoplanet visit, an Observation Notice is sent to all the participants alerting them to this new target, as
well as to which of the previous targets are still ones of interest.

As an incentive to amateur astronomers to participate in this project, if an observer’s data is included in the global fit that goes to the Hubble science team, any paper whose results rely on that data will have that amateur astronomer listed as a co-author.

7. The Future

The demand for amateur astronomer participation in exoplanet research will only increase in the next 1-3 years with the launch of two new space telescopes. For example, the Transiting Exoplanet Survey Satellite Telescope (TESS) is scheduled to be launched sometime in 2017-2018. TESS will conduct an all-sky survey for transiting exoplanets – a much broader survey than that conducted by Kepler. TESS will likely result in thousands of exoplanet candidates that need to be confirmed by ground-based observers. Efforts are already underway by the AAVSO to develop an archive for such observations. In 2018, the James Webb Space Telescope (JWST) is scheduled for launch. It will be observing exoplanet transits in the infrared and it too will require follow-up, ground-based observations.

Although the contributions of amateur astronomers to exoplanet research has been primarily through use of the transit method, high precision spectroscopes are predicted to soon be in the hands of amateur astronomers where they can now also conduct radial velocity studies using Doppler spectroscopy.

It is not clear that direct imaging of exoplanets will ever be possible for amateur astronomers. However, their use of modestly priced equipment to do spectral interferometry and resolve double stars at a distance of 0.5 arcseconds from each other (Ashcraft, 2016) was considered unlikely just a few years ago.

Finally, very recent observations by amateur astronomers have also begun to contribute to our understanding of so-called “exo-objects,” namely objects of an undetermined nature around distant stars. Such an example is WD 1145+017, which is suspected to be transited by a disintegrating asteroid. Its anomalous light curve was first detected using data from Kepler’s K2 mission (Vanderburg et al., 2015). Follow-up observations by amateur astronomers confirmed these strange, drifting light curve patterns (Rappaport et al., 2016). Amateur astronomer Mario Motta has more recently been making follow-up observations of WD 1145+017.

Figure 5 depicts such an observation by Motta on March 30, 2016 and Figure 6 shows a follow-up observation by Motta taken just a few days later on April 6, 2016. The change in the light curve during the predicted transit times shows the evolving the nature of this object.

8. Conclusion

The field of exoplanet observing by amateur astronomers has matured to the point where a set of Best Practices can be established, as described in this paper. Furthermore, it is now possible for an amateur astronomer to conduct exoplanet transits with a high degree of accuracy, using the same equipment used for deep sky imaging. Data resulting from such observations has already materially contributed to major professional studies, such as the HST collaboration. New space-based telescopes, such as TESS and JWST, will only increase the need for exoplanet follow-up observations by ground-based
amateur astronomers. Finally, even our understanding of so called “exo-objects” is enhanced by observations of amateur astronomers.

9. References


Abstract

The Panoptic Astronomical Networked OPtical observatory for Transiting Exoplanets Survey (PANOPTES) is a citizen science project which aims to build low cost, automated, robotic sky patrol camera systems which can be used to detect transiting exoplanets: planets orbiting other stars. The goal is to establish a worldwide network to image the nighttime celestial hemisphere 24/7/365. PANOPTES will search for exoplanets using the reduction in starlight caused when an exoplanet transits its host star. Individuals or groups can construct a PANOPTES station, tie it in the data reporting system, and contribute to the discovery of exoplanets across the large area of the sky not yet surveyed.

1. Introduction

The discovery of extra-solar planets, “exoplanets,” orbiting other stars has a longer history than the celebration in 2015 marking two decades after the first discovery of an exoplanet orbiting a normal star, 51 Pegasi b (Mayor and Queloz, 1995). That discovery, using radial velocity (RV) measurements, was the culmination of a number of reports of possible exoplanets, only a small handful of which proved correct. The first efforts used astrometric means to look for reflex motion of the host star on the plane of the sky (e.g., Bessel, communicated by Herschel in 1844) in response to the orbital motion of a companion. The discovery of Sirius B, the “Pup,” (by the Clarks, reported in Flammarion, 1877) validated this method, though the Pup was recognized to be a white dwarf star. Subsequent reports of astrometric evidence for exoplanets over the following century-plus proved false.

Even as the number of RV discoveries ramped up it was clear that selection effects limited the chances for finding and Earth-like planet, the Holy Grail. To better determine the size frequency and orbital statistics of exoplanets with fewer selection effects, what became NASA’s (2009) Kepler mission was proposed. It would monitor a section of the Milky Way for regularly repeated transits (the silhouetting of an exoplanet as it passes through the line of sight between star and telescope). It was believed, at the time that the precision photometry required had to be done outside Earth’s atmosphere. For small terrestrial planets in Earth’s size range, this is the case but it turns out not to be true for larger exoplanets. As early as 2004 the Trans-Atlantic Exoplanet Survey (TrES) discovered a 0.75 M_J (Jupiter mass) planet (Roi et al. 2004). Additional ground-based surveys have added to the list of transiting exoplanets, even as Kepler’s spectacular results have provided the desired statistics, including unexpected findings.

Characterization of discovered exoplanets is now getting more attention from the professional astronomical community. At any given time most of the sky is not surveyed because stars are too numerous and often too bright for professional telescope systems to devote time to. PANOPTES’ continuous full sky coverage will discover a large number of nearby exoplanets.

To address this, PANOPTES has been developed to survey of the night sky 24 hours/day, 365 days/year. Stations scattered around the world will provide the full (night) sky coverage, with sites sharing similar longitudes providing overlap and redundancy. These are valuable not only to reduce the effects of bad weather. Duplicate observations also validate observed events (of all types) captured by the cameras.

2. PANOPTES

PANOPTES consists of a dual camera system mounted on a German equatorial mount. The whole system is automated to image multiple fields on the sky over a whole night. Weather sensors determine the suitability of conditions for making observations. The system initiates and ends observations based on
the day-night cycle and weather. PANOPTES is designed to be left in the open, without a shelter of any kind. It has been in development and testing for more than 5 years, including operational observations on Mauna Loa, Hawai‘i.

PANOPTES is designed to be assembled by anyone with some skill with common tools. Parts lists and detailed instruction sets are available on-line. The total investment for an operational system is approximately $5,000 USD.

Some site preparation may add to this total. For example, cables for power and communication must be emplaced from a local source to the survey unit. It may be desirable to construct a foundation or concrete pad for the equatorial mount pier or tripod legs.

The PANOPTES system itself consists of three main components.

2.1 The Camera Box

The camera box (Figures 1 and 2) encloses the two cameras and lenses along with an electronics (arduino) board and a USB hub. One power and one signal (USB) cable come up through the mount to the camera box. The USB hub then splits the USB signal out to the arduino and to both cameras. The power goes to the electronics board which regulates it down to the appropriate voltage for the cameras. (http://www.projectpanoptes.org/hardware/camera_box.html)

2.2 The Mount and Tripod/Pier

The PANOPTES baseline uses an iOptron iEQ30 Pro mount to which weatherproofing is added. The mount can either sit on the standard tripod or on an optional PANOPTES designed pier (Figure 3) for more permanent installation. (http://www.projectpanoptes.org/hardware/pier.html)

2.3 The Control Box

The control computer, power supplies, and supporting electronics all sit in a weatherproof enclosure next to the mount (Figure 4). A PANOPTES unit is designed to be self-contained, requiring only input external power and internet connectivity. A UPS (uninterruptable power supply) circuit ensures that the unit can safely park in case of power outages and provides reliable operation through brownouts and short power failures. (http://www.projectpanoptes.org/hardware/control_box.html)

Fig. 1. Camera box ready for closure and operation.

Fig. 2. The sky end of the camera box, ready for imaging.
2.4 Software

All PANOPTES authored software is open source and available on github.

(https://github.com/panoptes)

The PANOPTES Observatory Control System (POCS; https://github.com/panoptes/POCS) has been written from the ground up using python.

The photometry algorithm which extracts accurate photometry from DSLR images which have a bayer filter array to encode color information was originally developed by Olivier Guyon and is described on his web page.

(http://www.naoj.org/staff/guyon/09allskysurvey.web/56photometry.web/content.html) The algorithm has been implemented in C, but will be wrapped in a data pipeline which will be written in python.

The acquired data are automatically analyzed and transferred to a cloud operated by Google.

2.5 Hardware & Software Status

The PANOPTES system is in the final stages of development. Instructions, parts lists, and software are already posted, though the electronics and software are not quite in final form. Materials for the complete system are not exotic and can be purchased from various suppliers now. Assembly can begin when the parts arrive.

The baseline design is meant to be assembled by anyone with a basic knowledge of electronics, computers, and astronomy who has a taste for building things/tinkering. There is no need to machine parts and only standard tools are required. All electronics are readily available and only need to be mounted and wired together.

Alternatively, negotiations underway with a commercial supplier are expected to yield the sale of complete kits ready for assembly and programming.

3. Program Status

After several prototypes, a demonstration unit has been installed at Imiloa. PANOPTES is ready to welcome additional survey systems into the program. The data acquired by a station, automatically sent to the Google cloud, can also be retained for individual research projects.

PANOPTES is designed to allow individual stations to personalize their system to their research needs. Individual upgrades/additional instrumentation are possible. Cooperative efforts between individuals, groups, and/or schools can spread the cost.

Join PANOPTES and contribute to cutting edge research on exoplanets and other wide sky coverage research opportunities.

http://projectpanoptes.org/
4. Acknowledgements

This publication was prepared by the Jet Propulsion Laboratory, California Institute of Technology, under a contract with the National Aeronautics and Space Administration.

Reference herein to any specific commercial product, process, or service by trade name, trademark, manufacturer, or otherwise, does not constitute or imply its endorsement by the United States Government or the Jet Propulsion Laboratory, California Institute of Technology.

5. References


A Bespoke Spectropolarimetrist

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Abstract

"Bespoke" refers to something custom or self-made. This paper continues the journey of an amateur learning spectro-polarimetry using an 18inch f3.5 reflector with homebuilt spectrometer modified to perform polarimetry, thus yielding the wavelength dependence of stellar polarization. Polarization of starlight occurs as the light passes through polarizing material in space near to the star or through a more distant, larger interstellar cloud. Polarization may occur by several different mechanisms, and may give insight into the presence of stellar winds or magnetic fields and into the polarizing material itself. Many stars show very substantial time variations in both polarization and direction. In the case of spectral Class B (hot) stars that show emission (rather than absorption) at the Ha line, the change of polarization with wavelength may give insight into where the emitting region is. I will show results from some 200 measurements on about 95 stars ranging down to mag 5. This is a new, challenging, and potentially fruitful endeavor for amateurs and for possible Pro-Am projects.

1. Introduction

Polarization of the light from a star can occur as the light passes through gaseous or particulate matter that is non-spherical and at least partially aligned. Alignment can occur from magnetic forces, stellar emissions (eg., solar wind), or other causes. The amount of polarization is usually small, generally below 1%, with most stars showing less than 0.1-0.2% polarization. The challenge is to measure the polarization (requiring very high signal to noise ratios) and to understand the physical basis for the polarization and whether it is in or near the star vs an interstellar cloud.

Most polarization measurements are made using broad band (filter) methods which improves S/N ratios; however, it is also feasible to convert a spectrometer to measure the polarization vs wavelength. This information can help in understanding the location and cause of the polarization.

Amateurs are increasingly exploring their role in performing spectroscopy; however, very few conduct polarization measurements. The purpose of this experiment has been for an amateur astronomer to explore the challenges of spectropolarimetry (S-P) and to identify whether amateurs can contribute usefully to the research in the field. Interestingly, a reading of the professional literature gives rise to great pessimism that an amateur can fruitfully do this work. The good news is that the work reported here shows that the amateur can in fact successfully accomplish the measurements.

In 2014, the author presented at SAS2014 the initial results of this experiment (SAS2014). Since that time, the equipment and methods have been improved, so this paper will only briefly review the equipment and methods used, but will explore in more detail the changes made since 2014 in the instrument and methods and will show samples from recent observations.

2. The Challenges of Polarimetry

S-P is definitely a challenging effort to undertake. It requires learning a new way of thinking about performing observations and of analyzing data to obtain a measurement. And then, new challenges in understanding what the measurements mean (the realm of astrophysics).

The actual S-P observation is similar to taking a spectrum, i.e., the target is focused on a slit and the resulting spectrum is recorded on a CCD camera. Because the measurements require a very high S/N (e.g., to measure a 0.1% polarization will require roughly 1,000,000 counts in a bin), the observation must be conducted with high precision over long exposures. The present work uses an 18inch f3.5 homebuilt Newtonian telescope, and a homebuilt f3.5 R=3000 spectrometer, so the equipment needed is not necessarily expensive or very difficult to build. Using this equipment, this work shows that it is
feasible to measure S-P down to about magnitude 5 stars (better than the magnitude 2-3 reported in 2014) with a total exposure of 2-3 hours.

Figure 1. Telescope/Spectrometer/Polarimeter

As described in the 2014 paper, the conversion of the spectrometer to perform polarimetry required two changes: addition of (a) a Half Wave Plate (HWP), and (b), a Wollaston prism.

The HWP (also known as a retarder) chosen is an inexpensive polymer film device mounted in a holder that allows rotation of the HWP to four specific relative angles (0, 22.5, 45, 67.5 deg). The HWP is installed just in front of the diagonal as shown in Fig 2. The purpose of the HWP is to successively rotate the plane of polarization a known amount to sample the various polarization angles in the starlight. After the HWP, the light passes normally through the slit and spectrometer.

Figure 2. Rotating HalfWavePlate Assembly below Secondary

The second change is to install a Wollaston prism just ahead of the CCD camera used to record a spectrum. The Wollaston prism (also known as an analyzer) is made of Calcite birefringent material and has the property of directing the incoming light rays to one of two output angles corresponding to whether the ray is polarized horizontally or vertically relative to the prism. Thus, the CCD camera will show two spectra, with each corresponding to the particular polarization direction.

Figure 3. Wollaston Prism Function

The actual observation requires four spectral images, one for each HWP angle (usually multiple exposures), with each image containing two spectra as shown in Fig 4.
Standard software is not available to analyze the spectra; however, it is entirely feasible to use fairly simple routines with standard amateur software to tailor the analysis process. For example, the present work uses custom Visual Basic routines controlling MaximDL to run the camera and HWP rotator and then to analyze the spectra images.

The analysis of the images requires application of darks and flats, digitizing the spectra, and compiling the results from the various multiple exposures. After pasting the data file into an Excel spreadsheet, Excel then analyzes the data and presents the results. Using these tools, the data from 3-4 stars during a night’s run (perhaps 125 images) can be analyzed in about 30 minutes.

Calculating the polarization from the spectral data follows the method laid out in the ING website (see References). With the spectra loaded, the spreadsheet computes a series of ratios among the eight spectra that cancel out intensity changes and other effects including differences in the polarization properties of the optical system for the two polarizations during the run. This process results in a graph of polarization vs wavelength such as shown in Figure 9 (note the vertical scale is 0-1%). One can also compute the polarization angle vs wavelength using a very similar approach (results not shown in this paper).

### 3. Initial Operation and Problems

At the time of the 2014 paper, it was already clear that amateur S-P measurements could be made, but problems were already apparent. These included:

- Limitation to only the brightest (mag 2-3) stars
- *Poor sharpness of the spectral images and a requirement for a custom lens following the Wollaston prism
- *Large artifacts in many of the polarization curves
- Guiding issues associated with the pellicle (think) pickoff guiding mirror.
- Unknown systematic errors (ie, one must prove that the system is giving the correct answers)

As an example of the very rough early test results, Fig. 6 shows Alkaid polarization as observed in 2014. These results seemed to indicate that the system was working to some extent, but had large amounts of noise and artifacts in the results. The quality of this result can be compared to current results in this paper, eg, in Fig 11.

Much of the poor sensitivity was traced to adverse effects of efforts made improve the quality of the spectral images. The specific problem is that the two spectra on each image (e.g., Figure 4) could not be brought to focus at the same time apparently due to the way light moved through the Wollaston prism. However, experiments showed that most of the defocus was in the Y-direction, ie, it did not affect spectral resolution. Using modified software to
accommodate the wider (poorly focussed) and narrower spectra allowed the correcting lens to be eliminated, in turn allowing the Wollaston to be moved closer to the camera, thus passing a greater fraction of the incoming light.

![Figure 7. Navi (mag 2, B0e) Polarization 11/22/15 before correction of HWP internal reflection](image)

While measurements in 2014 showed evidence of artifacts in the results, with improved sensitivity and precision these became overwhelming, as shown in Figure 7. This obvious artifact, a 22Å variation in the measured polarization, was present in most observations. Interestingly, even though the artifact dominated each polarization curve, the average value proved to be accurate and well behaved when compared to professional measurements, in this star about 0.5% with a dip at Ha (6563Å).

Initially, the artifact was thought to be an effect of the thin pellicle guiding mirror. However, replacing this mirror with a 3mm thick mirror did not affect the artifact (though it made guiding easier as the reflections from the front and rear surface were now well separated in the guiding camera).

Calculations showed that internal reflections in the HWP polymer film would have approximately a 20Å apparent wavelength. After finding a cooperative optical company, the HWP was coated on both sides with an anti-reflective film. This step completely solved the artifact problem (compare Figure 7 to Figure 9).

There are many sources of possible errors in polarization measurements. As just one example, limited statistics (i.e., poor S/N) can increase the "zero" setting of the instrument, as well as producing noise in the results. A minor 1% noise spike in a single bin in a spectrum can give rise to a large artifact in the polarization result.

Because of the compromises made in the design (e.g., using inexpensive polymer HWP vs expensive achromatic devices), there is potential for substantial errors due to the wide range of angles of the light entering the HWP and the Wollaston, and from the effects of the range of wavelengths passing through the system. Evaluating these kinds of errors requires extensive work using a variety of "standard" polarization stars. This extensive work has not been done; however, spot check comparisons of these results compared to professional results shows agreement to substantially better than 20% of the measurement (i.e., our measurement will agree with the professional result to better than one part in five).

Finally, many polarization measurements have shown polarizations at and below 0.1% so this appears to be the inherent lower limit for this instrument. This is substantially better (lower) than the literature suggested might be possible for an amateur instrument, especially using a fast Newtonian system.

4. Experience and Results

Using the modified equipment and hardware, observations have proceeded virtually every clear night since December 2015. Approximately 200 observations have been made covering approximately 95 stars. The system has proved stable and robust, with only minor changes to instrument and software settings over 3-4 months (in spite of a removal and remounting of the instrument during that time).

How were targets chosen? With no particular knowledge of high polarization stars, I started with stars having high brightness, i.e., brighter than mag. 3. The targets each night were chosen using TheSky set to show spectral class by color. Because hot (Class O and B) stars often show night to night spectral changes, and often are known to have substantial dust or gas, these were given preference. Even more, Class Be (emission) stars, known to have Ha emission regions associated with rings of extra-stellar material, would seem to offer the opportunity to detect whether the Ha emitting surface was behind or in front of the polarizing material (i.e., whether the polarizing material was associated with the star or was interstellar). In addition, a variety of stars of other spectral classes were also observed.

In all cases, the spectrometer was operated in the region 6300-6800 Å to include the Ha (6563Å) and He (6680Å) lines of the spectrum. Tests of the polarimeter have not yet been performed for other wavelength settings.

Using remote control, the observatory was opened, and telescope trained on the target. In most cases, a nearby bright star was observed to verify the various settings (e.g., pixel value for the center of the slit). The telescope was then trained on the target and the exposure series started, usually cycling through the HWP angles for 1-3 hours while taking five...
minute images. The data were downloaded and analyzed the next day as discussed above.

As noted, results are in the form of polarization curves for a wide variety of stars. In most cases, observations were repeated after several days or weeks to determine whether stars appeared to have substantial variation. Some of the most interesting stars have been measured more than a dozen times.

Results show that many stars of either high and low polarization have shown high stability, i.e., essentially no changes over weeks or months, showing that both stars and the instrument are stable. In contrast, many stars also show substantial variability even night to night. In the Orion region, there is a concentration of stars with high polarization, presumably showing the results of interstellar materials polarizing the light from many stars.

Following are several examples of typical results.

The first example describes results for several Class Be stars. Navi (SAO11482, HD5394) is a mag 2.1, B0e (emission) star. Fig. 8 shows the totalized spectrum (i.e., all the spectra added together) with the Hα emission line standing out. Each bin has approximately 2,000,000 counts. Fig 9 shows the measured polarization curve. The average over this wavelength span is about 0.6%, which was stable +/- 0.1% over all the measurements (>10). The variations shown are mostly noise related (even with 2 million counts of data). The decrease in polarization to 0.25% in the Hα region is real, i.e., it appears in each measurement.

A similar star (mag 4.2 B5e), SAO38980, is shown in Fig. 10. In this case, the polarization is 1.0%, but near Hα plunges to 0.08%. In both stars, broad band polarization is polarization imposed by the interstellar and stellar environments. However, the large drop at Hα shows that the Hα source is "above" the stellar (polarizing) atmosphere, and the minimum value (0.08%) is then a measure of the interstellar polarization in the direction of that star.

Of the 20 emission stars observed, 11 showed substantial decreases near Hα. In several stars, the Hα decrease also changed markedly with time. These distinctions would help distinguish the different stellar structures and/or viewing geometries.
The final example is cautionary note about working with measurements at the 0.1%. SAO110665 (HD16582 mag 4.1 B2) was first measured on 02/27/16 as shown in Fig. 12. It showed a rising polarization over the wavelength range with an average polarization value of about 2.5%. The next night showed a "level" polarization of about 0.2%, but the third night again showed a rising polarization. Nothing like this had been seen in any of the observations of this campaign. We continued observing every clear night (14) for the next month. The polarization remained low and level, until 03/21/16 which began another series of nights showing the high and rising polarization. The resulting tally of average polarization and slope are shown in Fig. 13. Unfortunately, in early April the star was setting immediately after sunset, ending the observing season. Bottom line: striking evidence for unusual phenomena.

At first, of course, the suspicion was an instrumental or observing error. However, no other observation showed this polarization behavior. Other stars in the same region of sky, even the same night, showed "normal" results. A careful examination of the spectral information available showed no correlation with any obvious feature. This star (like most bright stars) has been subject of many studies over many years; however, a literature search (Google) showed no unusual characteristics of this star and no references to such bizarre behavior.

At this point, it was only human nature to begin thinking that perhaps a real discovery had been made. As is usually the case, reality then intruded. While observing another star, the results also showed a sloping polarization. Again working the data, I found that the sloping polarization disappeared if the first spectral image taken was deleted from the data set (even though there was nothing notably wrong). Work is ongoing to identify the exact cause of the problem (presumably related to transient issues related to the guiding system). However, the experience was both humbling, and a reminder of how subtle errors can wreak havoc when attempting measurements at the sub-1% level. It also showed the importance of operating a new instrument in a "survey" mode both to gain experience and to identify potential errors.

5. Conclusion

The experiment of building an amateur spectropolarimeter, and learning how to use it, has been a clear success. The current data demonstrate that useful measurements can be made with reasonable accuracy. This opens the possibility of amateur contribution to the research literature, especially for purposes of monitoring changes in polarization of at least the brighter interesting stars. Attempts are continuing to establish relationships with professionals in the field to obtain guidance on methods and suitable observing targets.
6. Acknowledgement

I wish to thank Ted Dunham of Lowell observatory for his help and encouragement. Also, I acknowledge the wise advice of several of the optical vendors with whom I have worked.

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http://www.ing.iac.es/astronomy/observing/manuals/html_manuals/wht_instr/isis_hyper/subsubsection1.7.0.2.5.1.html#SECTION070251000000000000000

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714d0ec1a7ef%29/publications.html?page=26
HPOL Spectrometer at Univ. Wisconsin has excellent website with both instrumental and astronomical information. See
http://www.sal.wisc.edu/HPOL/
**Abstract**

Conventional wisdom says it should not be possible to measure stellar radial velocities with a useful degree of precision with a spectrograph having spectral resolution of 1000. This paper will demonstrate that with a combination of careful observational technique and the use of cross correlation it is possible to far exceed initial expectations. This is confirmed by reproducing the known radial velocity of a catalogued SB1 star with a precision of 5.2 km/s. To demonstrate the scientific potential of such a spectrograph, we use radial velocity measurements to confirm the binary nature and measure the orbital period and parameters of a suspected post common envelope binary.

8. **Introduction**

The precision with which a spectrograph can measure the line of sight or radial velocity (RV) of a star, denoted $v$, as a fraction of the velocity of light, $c$, is determined by the precision with which the wavelength of a spectral line can be measured, $\delta \lambda$, as a fraction of its wavelength, $\lambda$. The relationship between these quantities is known as the Doppler-Fizeau (more commonly just Doppler) effect and, assuming the velocity $v$ is much smaller than the speed of light, is expressed as $v/c = \delta \lambda/\lambda$.

As an example, if we measure the wavelength of the H\textbeta line (4861 Å) in the spectrum of a star with an uncertainty of 5 Å, then we would expect the resulting RV of the star based on that measurement to have an uncertainty of around 300 km/s.

The spectral resolving power, $R$, of a spectrograph is its ability to resolve two spectral lines at a wavelength $\lambda$ separated by a distance $\delta \lambda$ and is defined as $R = \lambda/\delta \lambda$. The full width at half maximum height (FWHM) of the Gaussian profile of an emission line is often taken as a measure of the resolvable separation at the wavelength of that line.

The spectral resolution of the LISA spectrometer (Shelyak 2016) is nominally 1000. This means that at the wavelength of the H\textbeta spectral line one would expect to be able to resolve lines with a separation of 5 Å. From the above, my expectation of the precision with which it should be able to measure stellar RVs was around 300 km/s. This assumes that the spectrum is recorded with a CCD camera which has at least two pixels per 5 Å of dispersion to satisfy the Nyquist criterion.

I set myself the challenge to see if this performance could be bettered, and if so, by how much.

9. **Equipment and spectral processing**

![Figure 1. LISA spectrograph on a C11 scope with SXVR-H694 and SXV-EX cameras.](image)

The equipment I use is shown in Figure 1. This is a LISA slit spectrograph which includes an Ar-Ne calibration lamp and a reflective slit for guiding. The LISA is attached to a C-11 scope operating at F/5.5 with a focal reducer and the equipment is carried by a pier-mounted Losmandy G-11. The equipment is
operated remotely to avoid any manual interaction during an observing run.

Spectral images are recorded with a Starlight Xpress SXVR-H694 CCD camera which has 4.54 micron pixels and the guiding camera is a SXV-EX. I normally use a 23 micron slit which matches the average seeing at my observing site. The dispersion is 1.8 Å/pixel, comfortably within the Nyquist criterion. The FWHM of lines in the Ar-Ne lamp spectra around Hβ is 2.8 pixels confirming the spectral resolution to be 5 Å.

Astroart (Astroart 2016) is used for camera control and slit guiding. The resulting FITS files are processed using ISIS (ISIS 2016) which performs bias, dark-field and hot pixel removal and flat-field correction. Spectra are wavelength calibrated using Ar-Ne lamp spectra taken before and after each set of target spectra. Correction for instrument and atmospheric response is applied using spectra of an A or B type star close to and at the same altitude as the target and recorded immediately before or after the target spectra.

10. Cross correlation

The first indication that this spectrograph might be capable of better performance than the above analysis would suggest comes from the wavelength calibration process. Wavelength calibration to convert pixel position in a spectral image to wavelength is achieved by fitting a 4th order polynomial to the measured positions of 15 lines of known wavelength distributed across the spectrum of the Ar-Ne lamp. The root mean squared (rms) residual of this fit is around 0.1 Å indicating that we may be able to determine the wavelength of a spectral line considerably better by analysing many spectral lines together than we can by measuring a single line.

This is the potential of the technique known as cross-correlation. This compares the position of many lines in one spectrum with the position of the same lines in another spectrum. It looks for the relative displacement of the two spectra which produces the best match between the spectra thereby maximising the cross-correlation function. The calculation is performed in \( \ln \lambda \) space where the displacement of lines is independent of their wavelength. The relative RV between the two spectra can then be calculated from the Doppler equation.

11. Observing V471 Tau

The incentive to investigate the RV potential of the LISA came from a discussion with Dr Danny Steeghs at Warwick University. He suggested that I should try to measure the RV of the secondary star of the post common envelope binary V471 Tau. This is a well-studied binary containing a cool K2V main sequence star and a hot white dwarf with an orbital period of about 0.5 d and known RV semi-amplitude of 160 km/s. This would be a good test of whether I could detect and measure a RV of this size.

I recorded spectra of V471 Tau on 9 nights and processed them in the usual way. For each spectrum I corrected for the heliocentric velocity of the earth round the Sun and computed the RV difference of the second and subsequent spectra relative to the spectrum on the first night using the cross-correlation function in ISIS. Like many professional tools such as IRAF, ISIS contains many of the functions that you will one day find you need for spectral processing.

Figure 2 shows the resulting RVs as measured, the same data phased on the known orbital period of V471 Tau with the best sinusoidal fit to the data, and an extract from a published paper showing the measured RV phase plot of the K star (Bois et al. 1988). It is clear that by using cross correlation I was able to detect the RV variation in V471 Tau and to estimate its amplitude with a reasonable degree of accuracy. The rms residual of my RV measurements compared to the sinusoidal fit is about 45 km/s, a factor of 7 improvement on my initial expectation.

As the object of this exercise was to measure the relative shift in spectral lines over time as the binary revolved around its local centre of mass rather than to determine its absolute RV, I did not calibrate the RV scale with respect to RV standard stars. Thus there is no particular significance to the zero RV level in this figure and others in the paper.

12. Improving observing technique

Given this encouraging start, the next step was to investigate how to improve the quality of RV measurements with the LISA. Experience showed that several factors are important if more precise results are to be obtained:

- the temperature of the LISA must be allowed to stabilise before starting to record spectra;
- the LISA must be carefully focused at the operating temperature;
- calibration lamp spectra must be taken frequently to ensure accurate wavelength calibration;
- telluric features should be removed before cross correlation;
- best results are obtained with stars of spectral types F to K which have many narrow metal lines.
To test this improved procedure using a known binary star, I searched the SB9 catalogue of spectroscopic binary orbits (Pourbaix et al. 2004) for a single-lined binary with RV semi-amplitude in the range 30-50 km/s, orbital period of a few days, spectral type F or G and located at a convenient position in the sky to observe. HD116514 is a single lined spectroscopic binary with period 5.939 days, RV semi-amplitude 37.82 km/s, eccentricity zero, V magnitude 9.3 and spectral type G4V.

Spectra were recorded in groups of ten bracketed before and after with lamp spectra and to improve signal to noise each group of ten was combined to produce a single wavelength calibrated spectrum.

Three such calibrated spectra were recorded on the first night and averaged to produce a reference spectrum. 23 similarly processed calibrated spectra were recorded on 11 further nights and RVs calculated for each spectrum relative to the reference spectrum using the cross-correlation function in ISIS including heliocentric velocity correction.

Figure 3 shows the resulting RVs as measured in the upper plot and in the lower plot the same data phased on the known orbital period of HD116514 with a sinusoidal curve based on the physical parameters of the binary. The error bars in the lower plot show the standard deviation for each night calculated from the spread of RV values measured that night. The mean of those standard deviations is 6.0 km/s. The rms difference between the measured RVs and the sinusoidal curve based on the physical parameters of the binary is 5.1 km/s. As before there is no significance to the zero point of the RV scale.

This exercise provided convincing evidence that with care and attention to observing technique it is possible to exceed the expected RV precision by a factor of more than 50.
13. A real application

Knowing that it is possible to measure stellar RVs with much higher precision than originally thought, what can usefully be done with this capability? Prof Boris Gaensicke and colleagues at Warwick University are interested in stars which show a UV excess in their spectra possibly due to the presence of a hot companion star. This indicates they might be post common envelope binaries which have not yet started to exchange mass but could one day evolve into cataclysmic variables.

One way to show that these stars are indeed binary systems is to detect RV changes in their spectra due to the motion of one or both stars around their common centre of mass. When we measure the relative RV of a star in a binary system, this is the product of the orbital RV of the star and the sine of the angle of inclination of the binary orbit. Thus binaries with a small orbital inclination to our line of sight will be harder to detect through RV measurement.

I started recording spectra of one of the stars on their list of candidates. This did not show any RV variation within my limit of detectability, either because it was not a binary or had a low orbital inclination so I moved on.

The next one I tried, TYC 555-445-1 with a V magnitude of 10.8, looked more promising. I quickly detected changes in its RV so recorded 42 calibrated spectra over 15 nights and subsequently added a further 3 a year later. As before I calculated the RV of each spectrum relative to the average spectrum measured on the first night. The mean of the standard deviations of the RVs for each night was 5.8 km/s. Figure 4 shows a typical spectrum of TYC 555-445-1. This is a close match to spectral type K2III.

As a check on the stability of my RV measurements, I also measured the relative RV of the RV standard star HIP15323 from the catalogue of radial velocity standard stars for Gaia published by Soubiran et al. (2013). This has V magnitude 6.4 and spectral type G0. The RVs measured for 20 spectra were essentially constant with a standard deviation of 3.5 km/s.

Frequency analysis of the RV measurements of TYC 555-445-1 as a function of time using the ANOVA (Schwarzenberg-Czerny 1996) method in Peranso (Paunzen & Vanmunster 2016) showed two peaks (Figure 5). Examination of the phase plots for both peaks clearly showed that the taller peak corresponding to a period of 7.59 days represented the correct orbital period.

Figure 5. ANOVA power spectrum of the RV measurements of TYC 555-445-1.

The upper plot of Figure 6 shows the measured RVs phased on a 7.59 day period while the lower plot shows the result of fitting the radial velocity profile of an eccentric binary to these data (Hilditch 2001). This fit gave the binary parameters listed in Table 1. The rms difference between the RV measurements and the fitted RV profile is 7.0 km/s.

Figure 6. Measured RVs of TYC 555-445-1 phased on the orbital period of 7.59 days (upper) and the fitted RV profile for an eccentric binary with the parameters given in Table 1 (lower).
Table 1. Parameters of an eccentric binary orbit fitted to the phased RV measurements.

<table>
<thead>
<tr>
<th>Parameter</th>
<th>Value</th>
</tr>
</thead>
<tbody>
<tr>
<td>Orbital period P</td>
<td>7.59 days</td>
</tr>
<tr>
<td>RV semi-amplitude K</td>
<td>26.4 km/s</td>
</tr>
<tr>
<td>Eccentricity e</td>
<td>0.08</td>
</tr>
<tr>
<td>Argument of periastron ω</td>
<td>303 deg</td>
</tr>
<tr>
<td>Projected semi-major axis asin i</td>
<td>3.94 R☉</td>
</tr>
</tbody>
</table>

14. Conclusion

It is clear that the LISA spectrograph is capable with careful observational technique of producing RV measurements considerably in excess of initial expectations for stars with F/G/K spectral types and it can therefore deliver results of real scientific value in the study of binary systems. RV measurements made with the LISA have confirmed the binary nature of the star TYC 555-445-1 and provided parameters of its binary orbit.

15. Acknowledgements

I am grateful to Prof Boris Gaensicke and his colleagues at Warwick University for their engagement with the amateur community which provides both guidance and motivation.

16. References


Crowd-Sourced Spectroscopy of Long Period Mira-Type Variables

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Abstract
Crowd-sourced observing campaigns provide frequent temporal sampling over long durations at lower cost than can be achieved with exclusively professional efforts. They have been used very successfully to construct wellsampled light curves for variable stars and inform the timing of professional observing efforts. Until recently, spectroscopy has not been a tool commonly available for use in crowd-sourced campaigns. But advances in commercial equipment and educational efforts have removed many of the traditional barriers to low-resolution spectroscopy with small telescopes. One type of target that could benefit from a crowd-sourced photometry and spectroscopy campaign is Mira-type variables with long periods (over 500 days). We report the preliminary results from a pilot study testing the efficacy of crowd-sourcing spectroscopic observations of those stars using small telescopes and filter wheel gratings.

1. Introduction

Mira-type variables are pulsating stars found on the HR-diagram near the asymptotic giant branch (AGB). They have amplitudes larger than 2.5 magnitudes at visual wavelengths and periods ranging from hundreds to thousands of days (Garrison, 1972). The stars probably range in mass from less than one to several stellar masses, making them highly evolved giants with significant mass-loss rates close to the transition to a planetary nebula (Willson, 2000). These objects are relevant to understanding the late stages of evolution for a vast majority of stars. Their mass loss also contributes significantly to the chemical enrichment of CNO and formation of dust in the interstellar medium.

Several Mira-type variables have changed their period, amplitude and/or spectral characteristics (i.e. LX Cyg, TT Cen, BH Cru). These changes are probably driven by the evolution of the star itself. Caught in the act, a transitioning Mira is a rare opportunity to directly observe rapid underrepresented stages of stellar evolution at key points in low-mass star evolution. These instances provide critical insight and data to refine stellar models. But often the changes have been noted after the fact and detailed observation has not been conducted during the transition (Uttenthaler, 2016).

The high subscription rate for national and university observatories makes it difficult to devote a substantial amount of professional telescope time to monitoring Miras. That effort is more efficiently conducted through crowd-sourcing involving skilled non-professionals. For more than a century the American Association of Variable Star Observers (AAVSO) have successfully crowd-sourced brightness monitoring of many long period variables including Miras. Advances over the last two decades in camera technology and low cost spectroscopy equipment for 6- to 10-inch telescopes make it possible to add spectroscopic monitoring to the existing crowd-sourcing capability.

Over the last year we have conducted a pilot study with small telescopes using CCD imaging and filter-wheel gratings to monitor Mira-type variables with the aim of determining what non-professionals may be able to contribute to long-term spectroscopic
monitoring. In particular we are interested in knowing what spectroscopic changes can be detected in a typical Mira over the course of its period as well as what the potential is for detecting the onset of key evolutionary changes in a Mira-type variable.

2. Expected Variations in Mira Spectra

The cool temperatures of the atmospheres of Mira variables (1500 – 3200 K) give rise to a spectrum similar to other late-type stars dominated by strong metal lines and molecular absorption bands (i.e. TiO). In all stars, the strength of the TiO bands typically increases at lower temperatures. As noted by Merrill (1962) the TiO bands in hotter Miras are somewhat weaker than would be expected in normal stars of the same spectral type. The average range of temperature variation for a Mira variable is about 500K over the course of its period (Garrison, 1972) so as they pulsate the bands are strongest at maximum brightness and weakest at minimum brightness.

Changes in mass loss rate will also cause some spectral variation over the star’s period (Willson, 2000). As a result, emission lines (particularly hydrogen) may be present at some times during the cycle (Garrison, 1972). “Line weakening” of metal lines such as Ca, Fe, and Cr (Merrill, et al. 1962) and variation of AlO absorption bands (Keenan, et al. 1969) may also occur as a result of emission driven by mass loss.

Mira-type variable stars can be sorted by their spectra into three types: M-type, C-type, and S-type. Each type is characterized by enhancements of specific elements, which imply different evolutionary states. As nuclear burning progresses in the star’s core different elements may be mixed with the surface material through “dredge-ups.” Those “dredge-ups” occur when the structure of the star is temporarily disrupted by a transition from one form of nuclear burning to another. A star that has transitioned from carbon fusion to oxygen fusion may exhibit a larger proportion of carbon and carbon-like molecules in its atmosphere. M-types are characterized by oxygen enhancement relative to carbon. C-types are more abundant in carbon than oxygen. S-types have about equal amounts of both carbon and oxygen and stronger ZrO, YO, and ScO molecular absorption (Merrill, et al. 1962).

3. Filter Wheel Gratings

Filter wheel gratings are diffraction gratings cut to a shape and size to fit into a standard slot in a filter wheel for a CCD imaging camera. They produce a slit-less first order spectrum of every star in the field. The spectral resolution depends on the line spacing ruled in the grating as well as the distance between the grating and the imaging plane. For most small telescope setups, the resulting spectral resolution is on the order of $R \sim 50 – 100$. To accurately map the wavelengths to the spectrum both the zero-order image and the first-order spectrum must simultaneously fit on the CCD chip. This provides a practical upper limit around $R \sim 200$ for most systems.

The throughput for these gratings is very good. Generally speaking, a usable spectrum can be obtained for any star that could be imaged with a S/N of 50 or better through a standard broad-band filter with the same telescope and CCD. Cost is likely not a barrier for most non-professionals who are already invested in CCD imaging. The cost of these gratings is typically a few hundred dollars, a small investment relative to the cost of a CCD camera and filters. They are also easy to operate because the images are taken the same as through any standard filter and only need a dark/bias calibration. Wavelength calibration takes some additional effort but only has to be done once if the distance between the filter wheel and imaging plane remains unchanged.

4. Observations

Stars for the study were selected from the AAVSO Long Period Variable (LPV) program. They were selected considering a convenient sky position for the observers, periods on the order of a few hundred days, and a minimum brightness that could be observed by small telescopes. See Table 1.

Two telescopes participated in this pilot study:

- In Pleasant Plains, Illinois the University of Illinois Springfield (UIS) Barber Observatory 20-inch reflecting telescope with Apogee U42 CCD camera using a back-illuminated E2V CCD42-40 chip. The plate scale for the imaging plane is 0.62 arcseconds/pixel. This setup used a Paton Hawksley Star Analyzer 200 grating yielding a first order spectrum with a dispersion of 1.328 nm/pixel in the imaging plane.

- Bill Rea operated an 80-mm Explore Scientific apochromatic refractor in Christchurch, New Zealand with an Atik 414E Mono CCD camera using a SONY ICX424AL front-illuminated chip. The plate scale in the imaging plane is 2.7 arcseconds/pixel. He used a Paton Hawksley
Star Analyzer 100 grating yielding a first order spectrum with a dispersion of 1.488 nm/pixel. For slit-less spectroscopy the image FWHM has a direct impact on the effective resolution of the spectrum. Assuming typical 3 arcsecond seeing, the effective spectral resolution unit for the UIS 20-inch was 6.4 nm. Rea’s 80-mm was 1.7 nm, almost four times better than the UIS setup.

The spectra in our study were extracted from bias-corrected and dark-subtracted images using the “projection” tool in SAOimage DS9. To correct for scattered light, a “sky background” strip extracted immediately next to the spectrum was subtracted from the target spectrum. Wavelength calibration was performed by applying a dispersion per pixel computed from observations of sharp emission lines in unresolved planetary nebula.

For more information, including a tutorial on wavelength calibration, see the web site: http://go.uis.edu/gratingspectra/.

The counts recorded in each spectrum varied depending on the telescope, exposure time, and brightness of the target. To aid comparison between spectra, their fluxes are normalized between 915 — 920 nm. These wavelengths are mostly continuum and relatively uninfluenced by telluric absorption or absorption normally present in Mira spectra. The normalization did not correct for the different spectral responses of the CCD chips.

### 5. Potential Issues with Crowd-Sourcing Spectral Monitoring of Mira’s

There are several issues that may present obstacles to successfully crowd-sourcing spectral monitoring on Mira-type variables with filter-wheel gratings. Those include: field crowding, spectral response of different CCDs, different spectra resolutions with different setups, focus of the images, and the practical faint limit for small telescopes. Foresight and planning may be able to overcome several of these.

#### 5.1 Overlap in Crowded Fields

Field crowding presents a challenge to all forms of slit-less spectroscopy. Without a slit the spectrum of the target can overlap the zero-order images or first-order spectra of other stars in the field. This presents a particular challenge to targets like Mira variables that are predominantly found in the galactic plane. When the target is much brighter than the other stars in the field this becomes less of an issue. The exposure time can be shortened for a brighter star so that the other stars in the field, which are much fainter, barely register in the image. However, a Mira variable near minimum can be faint, requiring deeper exposures that have greater potential for field crowding problems. Rotating the grating in the filter holder can change the orientation of the spectrum for the target but in very crowded fields this is not a practical solution. We encountered this issue with several potential targets, which were removed from the pilot study.

One solution could be to mask the grating in the filter wheel. A mask before the grating darkens the part of the field that the target’s spectrum falls in, eliminating any issue with overlap. Fast optical systems with the grating further from the imaging plane need to carefully account for the converging beam and the position of the grating when introducing a mask. For those systems, masking too much area around the target will cause a significant drop in the throughput for the first order spectrum. We suggest a mask that covers half the field, allowing the target to be positioned in the unmasked half of the field and the first order spectrum to fall in the masked portion.
5.2 Different Telescope/CCD/Grating Setups

One of the challenges of any crowd-sourcing effort is the difference in the equipment used by individuals who are contributing to the project. Rules and expectations can enforce uniformity, however too many regulations will discourage more people from joining the effort. Any successful crowd-sourcing endeavor needs to balance these competing factors.

It is unrealistic and burdensome to require all participants in a crowd-sourced spectroscopy project to have exactly the same telescope/camera/grating setup. We sought to gather information and suggest techniques and tools that can be used to overcome differences in the spectra contributed by different systems. The largest contributing factors to consider are differences in CCD response, differences in spectral resolution, and good focus of the images.

The two systems in our pilot study had some significant differences in those areas (see Section 4). We did not observe a common target with both telescopes, so we compare the images taken with either system of two M-type Mira variables with similar surface temperatures at about the same phase in their cycles.

![Figure 1. A comparison of spectra that should appear similar taken with the two systems used in this study.](image)

The differences in the appearance of the spectra in Figure 1 are due almost entirely to the differences in setup. We note that there is not a large difference between front-illuminated (UIS Barber) and back-illuminated CCDs (Bill Rea). Mira variables are very red. So even though the two CCDs have different blue response, their relative red response is similar (dictated by the atomic properties of silicon in the detector), so it is not a factor for comparing Mira variable spectra. We have also normalized the spectrum to continuum at 920 nm (see Section 4) where the relative response from both CCDs is similar.

There is a clear difference in Figure 1 due to spectral resolution. The Rea system has a higher effective spectra resolution (see Section 4) so the peaks and dips are sharper. Normal practice for comparing spectra of different resolutions is to blur the sharper spectrum to the lower resolution. Because that blurring degrades the better spectrum, we suggest that users of any archive generated by crowd-sourcing be made aware of the effects of differing resolution and provided with tools to degrade the higher resolution contributions as needed to make comparisons with lower resolution spectra.

It should be noted that neither setup has a high enough resolution to detect sharp emission features. Any analysis of Mira spectrum at this spectral resolution is restricted to focusing on the strength of absorption bands and not individual transitions.

It is important to encourage crowd-sourcing participants to get the best focus for their spectra and work under the best seeing conditions. Any problems with the focus will degrade the spectral resolution of their observations. Seeing and focus of the stellar images also contribute to the effective resolution. So contributors should be encouraged to avoid bad seeing conditions and frequently check their focus.

We also suggest that crowd-sourcing be restricted to only reflecting telescopes and apochromatic refractors. Rea’s own experience attempting to observe spectra with a two-element refractor showed that systems which have chromatic aberration will not be able to focus all parts of the spectrum to the same sharpness. This leads to a constant change in effective resolution across the spectrum and presents very significant challenges to comparing spectra from different telescopes.

5.3 Faint Limits

A successful crowd-sourced program should consider the ability of participants to observe the faintest targets. The light-curves of variables in the AAVSO database usually have many more observations of a variable at its brightest than when it is faintest. This can create problematic biases in analysis of the database if they are not recognized.

Our experience shows it is possible to get a spectrum with a filter-wheel grating for almost any star which can be imaged with a signal to noise acceptable for AAVSO CCD photometry. Field confusion and spectrum overlapping will be a greater concern in deeper exposures. But our suggestions for masking the grating in Section 5.1 potentially alleviate that issue.
6. What Was Detected

The resolution of spectra obtained with filter-wheel gratings is low. The spectra do not resolve or detect narrow emission or absorption features. The strong molecular absorption bands in Mira variables are detected as blends. The motivation for this pilot project was to determine if changes could be detected between different Mira types and also in a single Mira variable as it progresses through its cycle.

6.1 Differences in Type

Figure 2 shows that the differences between M, C, and S type Mira variables are easily detected at this resolution. The M type is dominated by TiO molecular bands. The C type shows almost no TiO but strong CN bands instead. The S type appears transitional between the M and C types including both TiO and CN bands.

Crowd-sourced Mira variable observations can monitor a large number of targets with regularity. Several Mira variables have changed type, but none has been caught in the act of changing. LX Cyg is an excellent illustrative example. Uttenthaler et al. (2016) report that sometime between 1975 and 2008 LX Cyg changed from type S to type C. Observations of the light curve in broad band filters did not immediately indicate the change and because of the sparse sampling of spectral observations its transition was not well recorded and only recognized after the fact. Crowd-sourced monitoring of the spectra has a greater chance of catching these transitions as they happen and providing a record of the changes and/or alerting professional astronomers to bring to bear additional resources.

6.2 Differences in Surface Temperature

Figure 3 shows the spectra of two M type Mira variables both at a similar phase in their cycle but with different spectral classifications (due to different surface temperature). Not only is there relatively more flux at bluer wavelength for the hotter star, but there are also, as expected, clear differences in the relative strengths of the different TiO molecular bands. It is difficult to flux calibrate spectra of this type so the difference in molecular bands is probably a more reliable indicator when considering data from several different telescopes.

6.3 Spectral Changes Through the Cycle

R Cen was observed continuously at regular intervals over more than half its cycle (2015 July 11 = phase 0.25 to 2016 April 5 = phase 0.77). The AAVSO light curve (Figure 4) shows that during that time it went through a local maximum at about 2015 Dec 01 = phase 0.53 as well as its minimum at about 2016 March 25 = phase 0.76.

Figure 2. Comparison of the spectra of three types of Mira variable.

Figure 3. A comparison of the spectra of two M type Mira variables at a similar phase in their cycle but with different spectral classes (surface temperatures).

Figure 4. Light curve for R Cen from the AAVSO.
(https://www.aavso.org/data/lcg)
Figure 5. Regularly sampled spectra of R Cen over the interval of phase 0.25 to phase 0.77 with flux normalized at 920 nm.

Figure 5 shows the normalized spectra of R Cen sampled at regular intervals during that time. The differences in relative flux are not dramatic but under careful examination they appear to correlate with the ups and downs of the R Cen light curve. An animation of these spectra is at: http://go.uis.edu/BarberObservatory/miras/.

We also have spectra of RZ Peg (S type Mira) and V Cam (M type Mira) near their maximum brightness and minimum brightness (Figure 6 & 7). Both show clear differences in their spectra but the differences in the spectrum of V Cam are more dramatic.

Because spectra at this resolution are sensitive to changes within a typical cycle it will be important to establish what is “normal” to differentiate them from evolutionary changes. The spectra around the cycle can also place limits on the mass loss and surface temperature fluctuations that are typical for each star.

7. Conclusions

As a result of our pilot study, we suggest the following guidelines for any crowd-sourced Mira variable spectroscopy project:

- Filter wheel grating spectra are sensitive to changes in:
  - Spectral class / Surface Temperature
  - Differences in Mira type
  - Different phases in a Mira variable’s cycle.
- Emission due to mass loss is difficult to detect at this spectral resolution.
- Pick targets that are not too faint for the telescopes used by your target audience.
- For Mira variables there is no significant concern about the difference in sensitivity between front-illuminated and back illuminated CCDs.
- Restrict crowd-sourcing efforts to reflectors and apochromatic refractors.
- Keep in mind the factors that influence effective spectral resolution. Encourage observing under good seeing and good focus.
- Archives of crowd-sourced spectra should provide tools to facilitate comparison of spectra with different resolutions.
- The continuum at 920 nm is a good point to normalize the flux of most Mira variables for comparison.
- We recommend masking half of the grating in order to get better results in crowded fields.

8. References


Abstract
The author describes how two commercial spectrographs, the SBIG DSS-7 and SBIG SGS, have been adapted for doing spectropolarimetry on a C-14 telescope. In the first case, to create an instrument capable of spanning the entire CCD range at a resolution of R~250. In the second, providing medium resolution (R~2500) and sufficient flux to clearly observe line depolarization in B[e] stars. The techniques described herein could easily be applied to other systems. The author discusses the design choices, operating procedures, and data reduction. A selection of observational results from both instruments are shown.

1. Introduction

The measurement of polarization provides insights into the nature of astronomical objects that are not otherwise obtainable.

Broadband and lower resolution spectroscopic polarimetry have revealed interstellar magnetism, stellar system asymmetries, asteroid mineralogy, and have enhanced exoplanet detection.

Higher resolution measurements of individual spectral lines and line groups have lead to better understanding of stellar atmospheres and have been critical in measurement of stellar magnetism.

As this paper will show, these techniques can be pursued with equipment familiar to amateur astronomers. While limited to brighter objects, time series observation provides a potential discovery space as yet largely untapped by professionals.

2. Polarimetric Resolution & Flux

In an optical spectrometer, the optical resolution is generally set by the projected width of the entrance slit upon the sensor and the dispersion of the light in terms of angstroms per sensor pixel. In the best case, every group of two adjacent pixels can be considered as an independent measurement. The quality of that measurement is the Signal to Noise ratio for energy collected in those two pixels. Since the brightness of spectral features varies greatly across the spectrum, much information is obtained with a S/N ratio of as low as 10. Hence collecting only a few hundred photons in each pixel can provide very useful science.

By comparison, the net polarization values for most sources are very small, and the variations that we wish to understand even smaller.

The 1 sigma error of a polarization measurement derived from n photons is $\frac{1}{\sqrt{n}}$. Thus to reach an error level of 0.1% requires recording 2,000,000 photons.

The 0.1% error level is a typical target for spectroscopic polarimetry. To achieve this, is is often necessary to combine adjacent pixels. The average number of pixels required, or the binning factor, multiplied by the dispersion per pixel, and divided into the working wavelength can be thought of as the polarimetric resolution.

Consider the case wherein the spectrometer slit is imaged onto two pixels with a dispersion of 1 A per pixel. The effective spectroscopic resolution is therefore approximately 3000 at H alpha. If each pixel receives an average of 10,000 photons, then the signal from 200 pixels must be binned together to provide a polarization measurement at the 0.1% error level. This in turn means that the polarimetric resolution at this wavelength is only about 30.

So to summarize, spectropolarimetry requires 5-7 magnitudes more signal than spectroscopy at the same resolution.

3. Dual Beam Polarimetry

The instruments described herein are based upon the principles of dual beam polarimetry. One of the comprehensive explanations of this technique for spectropolarimetry is found in Tinbergen(1).

Two optical components are commonly used to enable the measurement of polarization.
The first is a optical retarder or “waveplate” which has the ability to rotate the incoming plane of polarization.

Wave plates for the measurement of linear polarization are called half-wave plates because they shift the phase of the electric vector by one half of a wavelength. Such an optical element has the effect of rotating the plane of polarization by twice the angle that the wave plate itself is rotated.

This modulation allows the polarization of the incoming starlight to be correctly measured even after it has been modified by other optical elements in the system.

The second optical component is a beam splitting analyzer which divides the star beam into a double image so that the light in one image is orthogonally polarized with respect to light in the other. The ratio of intensities yields the degree of polarization across that particular optical axis.

For measurements in the visible and near IR region, these beam splitters can be constructed as calcite prisms in Savart or Wollaston forms. Savart plates, which are used in this work, split the beam into two parallel beams, with separation dependent upon the thickness of the crystal. Wollaston prisms create diverging beams, and so the final image separation depends upon the distance between the prism and the sensor.

Dual beam polarimetry depends upon both waveplate and analyzer elements. The analyzer splits the star image into a double image with orthogonal polarization. The ratio of the brightness of the two images reveals the degree of polarization along that axis. Since both images are derived from the same star, errors caused by scintillation and changes in sky brightness are eliminated.

By rotating the waveplate by 45 degrees the angle of polarization that reaches the analyzer is reversed. By dividing the first brightness ratio by the second, changes caused by the internal optics (mirrors, gratings, and pixel sensitivity) are cancelled out.

Repeating this process with the waveplate at 22.5 degree and 67.5 degree relative positions yields the polarization across a second axis. These two independent measurements are called Q and U.

The degree of polarization is simply the fraction of the incident flux that is polarized. For most astronomical objects this is no more than a few percent. The degree of polarization is calculated as the quadratic sum of the Q and U values.

The angle of the polarization ellipse is derived from the ratio U/Q and is reported in accordance with an IAU convention. This requires calibration by observing standard stars with known polarization. From this a starting “0” position for the waveplate can be determined and used for research observations.

4. Telescope System

The author’s observing system is built upon a Celestron 35cm Schmidt Cassegrain telescope with a motorized secondary mirror. This allows precision focusing while allowing the 18kg instrument stack to be fixed in place. An Astro-Physics AP1200 mount supports the assembly. The entire system is fully automated for data acquisition and reduction.

![Figure 1. C-14 instrument assembly. SPOL on left, instrument selector center, SGSPOL on right. The main camera is below the instrument selector, the waveplate carousel is in front of it.](image)

Light exiting the telescope proper passes through a waveplate carousel before entering a 4-way Optec instrument selector. From here it is directed into either the imaging camera or one of the two attached spectropolarimeters as seen in Figure 1.

These two instruments are called SPOL and SGSPOL. They operate at nominal resolutions of 250 and 2500 respectively.

4.1 The Wave Plate Carousel

Because reflections from right angle mirrors polarize the reflected light, the wave plate is normally placed before other optical elements.

In order to cover the entire CCD range at reasonable cost, two achromatic half wave plates are needed. This in turn creates the need to switch from one wave plate to another during observations. It is also necessary to provide high precision rotational positioning for each.

The wave plate carousel was custom made for this application by Optec Inc. At its core is a 5 position, 50mm filter wheel. For this application, 4 of the filter positions contain rotatable gear assemblies. In the center of each of these gear rings
is a 42 mm clear aperture in which a polymer wave plate can be mounted.

Once a particular wheel position has been selected, a geared stepper motor descends from the top of the device and engages the teeth surrounding that wheel position. The wave plate gear is then rotated until a home position is reached as defined by an embedded magnet and a Hall effect sensor.

Once that wave plate has been “homed” it can then be rotated to any of 5002 unique angular positions. Thus each of the required angular positions (0,22.4,45,67.5 degrees) is achieved with an error of less than 0.1 degrees.

Figure 2. Wave Plate Carousel. The large ring connects the instrument selector. Note the external gear rings. The actual wave plates were not yet inserted.

Three wave plates are currently in use. The first is a half wave retarder that is effective from 425 to 700nm. The second half wave retarder operates from 700 to 950nm. The third is a quarter wave retarder used for circular polarimetry. It also works in the 425-700nm band. These materials are products of Bolder Vision Optik, of Boulder Colorado.

5. SPOL – A low resolution, wide band spectropolarimeter

5.1 Design

The objective of this instrument was to produce high quality polarimetry over the full wavelength range of the CCD camera. It has been used successfully on stars down to about 9th magnitude.

This device was inspired by HPOL(2), a very successful instrument originally developed by the University of Wisconsin on a 0.9m telescope at Pine Bluff Observatory and now installed at the Ritter Observatory in Toledo, Ohio. A large database of HPOL observations are available online.

SPOL is an assembly of an SBIG DSS-7 spectrograph, an imaging camera, a Savart plate, a precision rotator, and a calibration lamp. It provides spectropolarimetry over a wide spectral range from 400nm to 950nm. The effective spectral and polarimetric resolution is around 250 on brighter objects.

5.2 Spectrograph

The DSS spectrograph is a folded design using a slit mask, an F10 collimator, a 600-line grating, and an F5 camera optic. While marketed for use with an ST402 camera over a 400-800nm range, work by John Menke(3) showed that it worked well with an ST8 camera. Using this extends the spectral range out to 1000nm.

What make this particular spectrograph so useful for this propose is that it contains two programmable actuators. One levers the slit mask in and out of the focal plane, the other tilts the grating to place either the first order or zero order image onto the sensor.

Using these actuators allows 4 useful imaging results:

1) An image of the sky (at reduced intensity) with the grating at the zero position.
2) An image of the slit mask illuminated by the calibration lamp.
3) A slit spectra of the calibration source or of a star using the slit.
4) A slitless spectra of a star that has been prepositioned at the normal slit position.

All four of these modes are used to enable spectropolarimetry.

The key design decision in any spectropolarimeter is the location of the beam splitter. It may be located before or at several points within the spectrometer. Each potential location has tradeoffs.
The DSS internal optics are fixed except for a focusing adjustment on the camera side.

Putting the Savart just before the CCD sensor allow the does normal use of the slit. Unfortunately, the fast f/5 camera optics aggravate a problem called Savart astigmatism which converts the round star image into an ellipse. This blurs the finest detail of the spectrum and tilts the spectral lines. This makes data extraction more difficult.

Hence the better of the available choices was to put the Savart external to the spectrograph and eliminate the use of the slit.

Removing the slit avoids the differences in the relative transmission of the two star images that result from seeing and guiding variations.

Since our polarization measurements are critically dependent on the ratio of brightness of these two images, slit induced differences are unacceptable.

Slitless operation greatly enhances the available flux. It adds at least one full magnitude to the instrument’s potential.

In the slitless configuration, however, the spectral resolution is dependent upon the seeing conditions at the time of the observation.

For the DSS working at a 4m nominal focal length, the standard 50 micron slit subtends 2.6 arc seconds. The projected slit results in an effective resolution of R~400.

The actual seeing-limited star images at this location vary from 3-4 arc seconds and so that true resolution varies from from the low 200’s to the mid 300’s.

Each target star is focused at the beginning of the observing process. This measures the size of the effective seeing disk, allowing the observation to be abandoned in very poor conditions.

The calcite Savart plate used in this system is 10mm square and 12mm thick. This provides a separation of 0.9 mm between the two images of a star. This is further divided in half by the spectrograph optics.

The Savart is mounted inside of a high precision rotator such that the end of the crystal is approximately 25mm from the nominal slit position.

5.3 Savart Rotator

The precision angular alignment of the Savart plate with respect to the sensor is critical to the successful operation of this instrument. To achieve this, the Savart is mounted in a high precision rotator. The positions of the two star images are aligned vertically as part of the setup process for each observation.

The need for this comes from the close relation between the resolution and dispersion of the instrument.

Most real spectral lines are less than an angstrom in width. The spectrograph optics broaden these lines to the size of the projected slit. Hence the entire variation from continuum to the spectral line center back to the continuum is compressed into two or three pixels. This amplitude variation can be very large. In strong emission lines the signal can vary by thousands of counts between the pixel with the peak value and those adjacent to it.

Now consider that fact that the polarization information in contained in the ratio of the flux between the pixels in the two parallel spectra. Assume, for this purpose that the starlight is not polarized. If the two spectra are in perfect wavelength to pixel co-alignment, then the two pixels representing the center of the emission line should have the same value (within statistical error.) Hence the ratio of values is unity and the measured polarization is zero.

But should the spectra be even fractionally displaced in wavelength from one another, then the flux of the peak of the displaced emission line will be partially allocated to an adjacent pixel. The ratios, therefore, will no longer be unity, but will be large values because of the acute spectral slopes. Thus the computed polarizations graph will show large artifacts in the areas with strong spectral features.

The spectrograph is bench aligned so that the dispersion accurately follows the pixel rows. The rotator is then used to place the two star images centroids into the same column with sub pixel accuracy. This allows the spectra to be co-divided without serious artifacts.

5.4 Calibration Spectra

A small segment of plastic optical guide is used to bring light from an externally mounted calibration source into the light path. A simple fluorescent lamp provides easily recognized lines from near 4000 to 9100. These are measured and recorded as a calibration table for each observation.

5.5 Imaging Camera

This instrument has been used with both ST8 and ST9 series cameras. The wider sensor of the ST8 and smaller pixels make it a better choice, but both produce good results.
5.6 SPOL Observing Process

1) The scope is centered on a sky position about 10 arc minutes from the target star.
2) The light is switched to the spectrograph by rotating the mirror of the instrument selector. A series of 3 “sky dark” frames and a bias frame are recorded. These will be median combined and scaled for dark subtraction in the spectral images.
3) Using the main camera, the star is positioned so it will be seen by the spectrograph in imaging mode. The light beam is then switched back to the spectrograph.
4) The calibration lamp is turned on.
5) An image is taken with the slit mask in place, and the grating in the zero order position. Software locates the position of the center of the smallest slit.
6) The grating is moved into its normal position and an image is taken of the spectra of the calibration lamp. From this image a wavelength calibration table is prepared.
7) The calibration lamp is turned off.
8) The spectrograph is returned to imaging mode and the star pair is moved near to the slit center position.
9) The star is focused.
10) The star pair is aligned to the sensor column direction using the rotator.
11) The waveplate carousel is set to the first waveplate and the preset 0 angle position for that waveplate.
12) The star pair is moved to the slit center position derived above.
13) The grating is moved into the dispersion position, and a spectrum test exposure is taken. From this, an exposure is calculated for all the data frames to follow.
14) A set of 4 spectrum images are taken and dark subtracted using the sky dark obtained earlier from the nearby sky background. The wave plate is moved from 0 to 22.5 to 45 to 67.5 degrees for these exposures. This set of 4 images is called a “data set”. After each data set, the star pair is realigned to the initial position. The relative position of the O₂ absorption minima is measured in each image for use in aligning the spectra pairs for calculation.
15) Additional data sets are collected if required. If both V and I band polarimetry is required, then the waveplate carousel will be switched to the I band waveplate and one or more additional data sets will be collected.
16) Keywords are created for each image that identify the data set and waveplate. These are recorded in an observation log file with the actual file name of the image. That log file is then used by the calculation process.

6. SGSPOL – A medium resolution line profile spectropolarimeter.

6.1 Design

This instrument was assembled for the purpose of observing the polarization phenomena associated with emission lines in B[e] stars.

In addition to the shared telescope optics, waveplate carousel and instrument selector, there are three components: a Savart positioner, a modified SGS spectrograph, and an SBIG ST7XME camera which acts as positioner, auto guider, and spectrum imager.

6.2 Savart Positioner

The initial instrument configuration placed the Savart plate into a rotator, analogous to the SPOL design. After some experience it was realized that the rotator was not needed.

Unlike the SPOL, where the entire Ha line spans 3 pixels, the SGSPOL emission lines can span as many as 100 pixels. Hence simple software interpolation provides excellent co-alignment without artifacts.

The rotator was replaced with a filter wheel. The Savart was placed in one position and the other positions left empty. This allowed the instrument to be switched from spectrometry to spectropolarimetry under program control.

Figure 3. The SGSPOL with Savart in place.
6.3 Spectrograph

This SBIG SGS spectrograph contains 600 and 1800 line gratings, the latter providing a dispersion of 0.29 Å/pixel at the Hydrogen alpha wavelength.

The optical system is capable of imaging an 18 micron slit clearly onto two 9 micron pixels so as to provide an effective resolution near 10,000.

In this application the slit has been replaced by a thin glass window. Thus the resolution is determined by the seeing disk. Typical 3-4 arc second Reno, NV seeing produces an effective spectral resolution of R~2200 to R~2800.

Removing the slit not only ensures that the flux ratio in the double star image is preserved, but it also increases the transmitted flux by a factor of 6 or more. This is essential in order to have practical exposure times.

As a final refinement, the original all glass slit disk was replaced with an assembly consisting of a 25-micron reflective slit on a square glass substrate abutted to a coated glass square. This assembly was mounted on a thin metal plate fabricated to mount in the original slit position.

A star image placed on the slit plate creates a projected spectrum near the top of the CCD frame, while one on the glass is imaged about half way down.

This allows the device to function as a slit spectrograph, a slitless spectrograph, or a slitless spectropolarimeter by choosing the star guide point and filter wheel selection. It also enables reference lines from a neon calibration source to be measured, a feature not available without a slit.

6.4 SGSPOL Observing Process

a) The target star is positioned by the main camera to a preset image position and the beam is switched to the SGSPOL arm.
b) The star image is focused on the reflective part of the slit plane.
c) The star image is moved to a predefined spot on the glass side of the slit assembly. At this point it appears as a double image due to the reflection from both the front and back of the glass.
d) The Savart crystal is selected in the filter wheel. This now creates a 4-star pattern in the guide camera.
e) The focus is adjusted by a fixed offset to compensate for the refraction of the calcite.
f) After a final position adjustment, the SGS auto-guiding function is engaged on one of the star images.
g) The Visual band wave plate is selected in the carousel and set to the 0 position for this instrument.
h) One or more data sets are taken. Each data set consists of 4 exposures using the 0, 22.5, 45, and 67.5 positions of the waveplate. The typical exposure time used is between 5 and 15 minutes per frame. Keywords for frame are recorded with the actual image file name in the observation log file.
i) When all datasets are complete, the auto guiding is stopped and the light beam is switched back to the main camera.
j) A previously recorded dark frame is subtracted from each image.

7. Post Processing

When the observation is complete, the observation log lists the image files obtained for each data set.

The post process consists of 1) spectral extraction, 2) data integration, 3) data binning, and 4) polarization calculation and reporting.

7.1 Spectral Extraction.

a) A small set of columns at the image center are scanned to find the row position of the center of each spectra. This is done simply by finding the maxima of the sum of two pixels that are separated by the spacing characteristic of the Savart plate for that instrument.
b) Using this row as a staring point, vertical scans are made at the 25% and 75% image width columns to find the maxima at these points.
c) Using these two pair of row and column coordinates, the approximate center of the spectral maximum can be computed for each column in the image.
d) A subset of column pixel values, typically 17, is collected centered around the computed row position for that column.
e) The pixel values within his subset are sorted by value. The 7 highest are combined to be the measured value for that column. This process tolerates errors in the central position, line curvature across the image, and avoids the need to interpolate pixel values.
f) Once values have been obtained for each pixel, a wavelength is calculated using the
calibration table created during the observation.

g) For the SPOL data, a flux adjustment is applied. This uses a table of correction factors derived from the spectrum of an incandescent lamp. The flux adjustment is not used for the SGS because of the small spectral range.

h) The output of this process is a text file listing, for each column in the spectrum, the central wavelength of that column, the sky corrected signal count for that column, and an approximately flux corrected value for that wavelength.

i) A graph is created from the extracted values plotting both instrumental and flux adjusted values. These graphs are primarily used for visual quality control.

7.2 Data Integration

Ideally each observation would produce 4 pairs of spectra with the identical wavelength scales and very large signal counts each pixel.

The actual results obtained by the instruments described in this work are not ideal. First, the individual CCD pixel wells are limited to ~70K photons. Moreover, in the SPOL case, the spectrum spans a wide range of wavelengths, but the exposure time must be limited to avoid saturation at any wavelength.

Hence to get a useful level of signal across the whole spectrum, multiple data sets must be taken and then combined.

Two wave plates are used to cover the entire spectrum. Hence the multiple datasets in each band must be sliced and joined to create a continuous spectrum.

The data integration process builds a table consisting of 9 columns and as many rows as there are pixels.

The first column is loaded with the wavelength vector from the top spectrum from the first image of the first V band dataset. The second column is loaded with the intensity vector for that spectra. The next column is loaded with the intensity vector of the second spectrum from that same image, shifted and interpolated as needed so as to match the wavelength vectors of the second spectra with the first.

This is now repeated from for the three remaining image frames of the dataset.

In the SPOL case, the two spectra derived from a single image are assumed to have identical wavelength vectors, so no interpolation is done between the top and bottom vectors of each frame. However, because some drift may have occurred between images, the wavelength scale may be displaced from one frame to another required interpolation for proper alignment. The displacement was determined by finding the prominent O$_2$ absorption minima during extraction and recording it in the observation log file.

In the SGSPOL case, the top and bottom spectra in each image always have an offset with respect to each other. These offsets are used for all the images in the observation. This works because guiding is active for the entire observation, keeping the spectra in consistent positions in all frames.

Subsequent V band datasets are now aligned to the wavelength scale of the first dataset and the pixel values are summed.

If I band datasets are available, the system sets a split point at the first column above 7000Å. The I band data replaces the V data for all subsequent pixels. These values are again interpolated to be consistent with the common wavelength scale. Additional I band data sets are added until all have been consumed.

At this point the data integration is complete. All values from all pixels in all of the spectra have been summed in 8 columns of intensity values aligned to a common wavelength scale.

7.3 Binning

Polarization counts are subject to Poisson statistics, the standard error in an intensity measurement is equal to the square root of the photon count.

Since each spectral column has a unique signal level, it follows that there is a different level of error in the polarization derived for each wavelength. This would make the interpretation of the results very difficult.

The solution to this is simply to group columns together in such a way as to get a similar total signal level for each group. This process is called binning.

For the SPOL instrument binning is done from left to right. A pre-specified error level determines the number of signal counts needed for each independent polarization calculation.

For example, an error level of 0.1% sets a requirement for 2x10$^6$ counts per bin.

The binning process simply adds the values from all eight of the data columns for the first pixel. If this is less than the requirement, the values from the next pixel are added to the running total. This is repeated until the end of the spectra is reached or the required count achieved. In this way a range of pixels has been selected. Next the eight spectral values from each pixel in the group are individually co-added.
These 8 composite values are now used to determine the polarization results for the wavelength range of the selected pixels. The binning process restarts with the next unused pixel and is repeated to the end of the spectrum. In this way a series of polarization results are generated with similar error statistics. The wavelength range of each result is simply a function of the signal level of the data. Hence the effective polarimetric resolution varies across the spectra.

For the SGSPOL, the entire spectrum represents a very narrow range of wavelengths (~200Å). In order to clearly report on the polarization behavior at H-Alpha, it is most useful to bin outward from the line center. For maximum sensitivity, the entire line profile width is used as a binning group and the full continuum on either side used for comparison. In the brightest objects, finer binning has been used to measure variation across the line profile.

7.4 Calculation and Reporting

The eight intensity values derived from each bin are used to calculate Q and U polarization. Then these are used of calculate the percent polarization P% and the angle of polarization, Theta. See Tinbergen(1).

The results are graphed in a format called a Triplot. In a Triplot, the upper curve shows the spectral flux. The middle shows the polarization. The lower shows the angle of polarization. All three use a common wavelength scale. Each data bin appears as flat horizontal segment in the graph.

8. SPOL Observations

Strong stellar polarization was discovered by Hall and Hiltner in 1948. Soon thereafter it was realized that this polarization was being produced by interstellar dust aligned by magnetic fields. In a 1973 paper Serkowski, et al. provide a mathematical formulation for the spectral distribution of this polarization. The observations below clearly show the characteristic shape of this polarization curve. A comprehensive summary of this phenomena is included in the 1974 Gehrels compendium (4)

8.1 HD 154445

This is a 5.45 magnitude star of B1V spectra type. It is standard star for used for polarization calibration. It polarization curve shows the typical appearance of polarization caused by the interstellar medium.

8.2 HD 161056

A magnitude 6.5 star with a B1.5V spectra. This is another star used as a polarization standard which displays a high level of interstellar polarization.
8.4 Epsilon Aurigae (SAO 39955)

The author has considerable experience in monitoring the polarization of this fascinating star system during and after the 2009-2010 eclipse. There is interstellar polarization of just under 2% and an intrinsic variation of between 0.1% and 0.2%. It is hoped that time series spectropolarimetry will lead to some additional understanding by revealing the polarization spectra of the intrinsic component.

To this end the author is using nightly broadband measurements to trigger SPOL observations of the high and low states.

9. SGSPOL Observations

The SGSPOL was implemented largely in response to the author’s interest in the fascinating work of Dr. Rene Oudmaijer of the University of Leeds.

His comprehensive review of the polarization behavior in circumstellar disks can be found in arXiv reference(5).

Most stars do not show intrinsic polarization, because the polarization that is produced is symmetric and so cancels out.

But when a star has an accretion disk structure, the disk breaks the symmetry, and a low level of net polarization can be created.

Unfortunately, the same polarization can be created from interstellar dust effects.

The hydrogen emission lines generated in the disk are not efficiently polarized and so provide a way of separating the interstellar from the local polarization.

This allows the intrinsic polarization of the circumstellar disk to be studied.

Line profile spectropolarimetry provides information on spacial scales that cannot be directly observed.

This work requires 0.1% polarimetry and good spectral resolution which SGSPOL can provide on
bright objects. It is likely that this is the smallest instrument that has used for these observations.

9.1 HD 5394 (Gamma Cas)

HD 5394 is 2.39 magnitude Be star showing strong emission at Hydrogen alpha. Because this emission occurs in a circumstellar disk, it is less polarized that the light generated nearer the photosphere.

9.2 Beta Lyrae

Beta Lyrae is a 3.42 magnitude eclipsing multiple star system of spectral class B8.5 showing strong H-alpha emission, cyclical changes in the line profile, and variations in polarization.

10. Instrumental Performance

The relative efficiency of these instruments was determined by comparing the spectral intensities with the V band signal of the broadband polarimeter installed in the same system. After correcting for differences in camera quantum efficiency and dispersion, it was found that the relative efficiency of SPOL is about 32% and the SGSPOL is about 8%.

11. Conclusions

Small telescopes using commercially available spectrographs can do useful spectropolarimetry. Slitless operation opens up a substantial number of targets, and time-series spectropolarimetric studies offer the potential for discovery. Automation is essential for the routine use of these relatively complex observational and analysis procedures.

12. Acknowledgements

The author wants to express his thanks to Dr. Robert Stencel of the University of Denver for helping him get started in scientific polarimetry. Thanks also to Jeff Dickerson and his colleagues at Optec, Inc. for developing the custom hardware the system is based upon. And to Roberto Gonzalez of Bolder Vision Optik for supplying the achromatic polymer wave plates that make these instruments possible.
The author wishes to acknowledge his extensive use of the HPOL spectropolarimetric database for instrument development and calibration.

This research has made use of the SIMBAD database, operated at CDS, Strasbourg, France.

And the author acknowledges with thanks the variable star observations from the AAVSO International Database contributed by observers worldwide and used in this research.

13. References


Measuring Starlight Deflection during the 2017 Eclipse: Repeating the Experiment that made Einstein Famous

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Abstract

In 1919, astronomers performed an experiment during a solar eclipse, attempting to measure the deflection of stars near the sun, in order to verify Einstein’s theory of general relativity. The experiment was very difficult and the results were marginal, but the success made Albert Einstein famous around the world. Astronomers last repeated the experiment in 1973, achieving an error of 11%. In 2017, using amateur equipment and modern technology, I plan to repeat the experiment and achieve a 1% error. The best available star catalog will be used for star positions. Corrections for optical distortion and atmospheric refraction are better than 0.01 arcsec. During totality, I expect 7 or 8 measurable stars down to magnitude 9.5, based on analysis of previous eclipse measurements taken by amateurs. Reference images, taken near the sun during totality, will be used for precise calibration. Preliminary test runs performed during twilight in April 2016 and April 2017 can accurately simulate the sky conditions during totality, providing an accurate estimate of the final uncertainty.

1. Introduction

Albert Einstein published his theory of general relativity in 1915, and soon made the startling prediction that the sun’s gravity would deflect light twice as much as Newtonian physics indicated. He calculated that the deflection of light from a star appearing just at the edge of the sun would be about 1.7 arcseconds, making it appear slightly shifted. This is shown diagrammatically in Figure 1.

Sir Arthur Eddington proposed that this miniscule deflection could be measured during a solar eclipse, and several expeditions were attempted before success at the 1919 eclipse. Those early measurements were not very accurate, however [Will 2015; Kennefick 2009], and later measurements had only a slight improvement [Will 2010, Will 2014; Friesen 2011] as shown in Figure 2.

The most recent attempt [Brune 1976; Jones 1976], organized by the University of Texas, required moving 6 tons of equipment to Africa and leaving it there for 6 months in a guarded shed. The telescope was a 200 mm aperture refractor with a 2.1 meter focal length. The images were recorded on 12” glass plates. This was a heroic experiment, but only achieved an 11% uncertainty; with today’s technology, a better result should be obtained with a much smaller effort. Since radio telescope observations have made ultra-precise measurements of deflection down to 0.0002 arcsec, this new experiment is simply a celebration of the original experiment.

In order to measure the deflection of light by the sun’s gravity, an experiment needs to be set up very carefully. Basically, three things are needed:

1. Get good images with some bright stars,
2. Determine where the stars should be in the sky, and
3. Accurately measure the location of the stars in the images.

Calculating the differences between the expected and measured positions gives the gravitational deflection.
This requires a very good telescope, a very good camera, and very good experimental design. To measure the difference between the expected and measured location of the stars in the image requires a very good star catalog with several small, but important corrections. Finally, I need to measure the locations of the stars in the images with very small uncertainties, ultimately reaching an average error of only 0.01 arcsec. With today’s technology, CCD cameras can replace glass plates, image processing software can replace scanning micro-densitometers, and satellite-measured catalogs eliminate the problem of measuring the stars six months before or after the eclipse. This makes the experiment much simpler and should lead to much higher precision.

2. Equipment and Experiment Design

This experiment is made feasible by the availability of superb, commercially-available amateur astronomy equipment. Some of these items were not even dreamed of in 1973, and have been vastly improved even since the 2006 eclipse. After carefully analyzing all of the requirements for this experiment and comparing those requirements with a wide variety of telescopes and cameras, I selected what I believe is the optimum combination.

2.1 Telescope

The ideal telescope is the Tele Vue NP101is refractor, shown in Figure 3 [www.televue.com]. This telescope is small enough to be portable, but its 101 mm aperture is large enough to capture 10th magnitude stars with 1 second exposures. Its diffraction limit is only 1.3 arcsec at 630 nm, much smaller than the 2.5 arcsecond daytime seeing I expect to encounter. The short focal length of only 540 mm allows a wide field of view with a medium format camera.

2.2 Camera

Once the telescope was chosen, a wide variety of astronomical cameras were reviewed against the experimental requirements. The clear winner was the monochrome Microline 8051 CCD camera from Finger Lakes Instrumentation [www.flicamera.com] shown in Figure 4. An interline CCD sensor was required so that no mechanical shutter was needed. A large format sensor was desirable, but too many pixels would require too much time to digitize the images. Since totality only lasts 140 seconds, this speed is critical. A larger format camera might image more stars, but the number of stars per second is optimized for the 8 MPixel 8051 model. This sensor’s pixels are only 5.5 microns wide, a perfect match for the NP101is telescope focal length, giving 2.1 arcseconds per pixel (resulting in a 2° wide field of view). While this pixel size has a moderate size full-well capacity, the resolution requirement is more important. It turns out that the stars in the neighborhood of the sun during totality have a small range of magnitudes (7.4 to 9.5), so the dynamic range is
not too important. Exposures will be bracketed to make sure at least one-half of the frames are useable, so this also mitigates the dynamic range concern.

This ML8051 camera digitizes at 12 MHz, so digitizing a full frame takes only 0.7 seconds. If the exposures range from 0.2 seconds to 1 second, then over 100 images can be saved during this short eclipse.

The camera will be cooled to reduce readout noise, but since it will be operated during the daytime, the focal plane temperature might not fall much below 0° C. The background signal noise will probably overwhelm the readout noise or dark current noise, but temperature stability during the eclipse and the calibration phase is very important. The built-in fan will be operated at a reduced speed to minimize any vibration.

The camera is mounted with a T-mount flange that mates directly with the 2.4” diameter Tele Vue focuser. This makes the camera mounting very stable, further reducing any camera-telescope drift.

2.3 Mount

The telescope mount needs to be set up and polar aligned before the eclipse, and hopefully, the weather will cooperate so this can be done the previous night. A portable mount is required that can handle the NP101is and the FLI ML8051 camera. The Software Bisque [www.bisque.com] MyT Paramount, shown in Figure 5 on its standard field tripod, meets this requirement. The particular mount to be used in this experiment has a periodic error correction only a few arcseconds, and with permanent PEC, the tracking error measured less than one arcsecond. Setting up this mount in the daytime might be required, but that can be done to less than one degree polar alignment error using the built-in scales and using a hand-held GPS to determine true north. This amount of polar error creates only about 1/4 arcsecond tracking error per 1 second exposure, much smaller than the errors due to seeing or diffraction. Tracking error can be ignored, but a nighttime polar alignment will make eclipse-day automation less risky.

2.4 Sky brightness data near the sun

Since this eclipse has only 140 seconds of totality, there is no time to experiment with different exposures. Fortunately, there is one example of calibrated brightness near the sun [Viladrich 2016]. I used this data to predict what to expect during the eclipse.

An important pre-eclipse issue is to determine which stars can be seen during totality. After looking at hundreds of eclipse photos from dozens of posted web sites, I found only one eclipse chaser who used an astronomical camera to image during totality. Fortunately, he also took dark frames, used a monochrome sensor, and saved his files in FITS format. Christian Viladrich of France used an SBIG STL-11000 camera during the March 2006 eclipse from Egypt, using a similar telescope. His exposures were only 5 msec long, since he wanted to image the...
inner part of the corona. I first stretched his images by factors of 40 and 200 to get simulated images of 200 msec and 1 second. Ignoring CCD blooming, I then estimated the background brightness levels near the sun, and calculated how bright a star would be visible. The calculations included the noise due to the background light over a small number of pixels and then estimated the signal-to-noise ratio (SNR) for the different stars in the field of view. Since I need to accurately measure the centroid of the star, I required the calculated centroid error be less than 0.1 arcsec, leading to a minimum SNR of 16. The results are illustrated in the next figure. While more than 60 stars brighter than magnitude 12 are in the field of view, I hope to get good measurements from 7 or 8 stars with a limiting magnitude of 9.5. Note that the sun is offset from center in order to maximize the number of measurable stars.

The two stars that appear closest to the sun will have a larger gravitational deflection, about 1.2 arc-seconds. However, the corona will be variable and the sloping background might make those measurements unreliable. The average deflection of the dimmer stars that fall near the edge of the field of view is only about 0.4 arcsec, but appear on a flatter background. Since the FLI Microline camera downloads images in less than one second and no mechanical shutter is required, the plan is to take as many images as possible with both exposure durations.

2.5 Zenith sky brightness

During the partial eclipse phases before and after totality, there have been a few measurements of overall or zenith sky brightness [Sharp 1971; Silverman 1975; Möllmann 2006; Zainuddin 2009; Strickling 2016]. If the sky is dark enough during this partial phase, then some star fields near the sun could be imaged to provide the necessary calibration data. Dozens of stars in each image are needed to measure the plate scale and the optical axis to sufficient accuracy. Unfortunately, this is not possible until totality, based on the following analysis.

Figure 7 shows the sky brightness during several eclipses, measured using all-sky photometers. There is a wide variation in brightness, but the conclusions don’t change. The sky brightness has been converted to a visible star magnitude, normalized to magnitude 9.5 for the time during totality. Even 30 seconds from totality, the stars need to be magnitude 6 to be visible. The Pleiades cluster (needing about 20 seconds slewing time to reach) might show only 10 stars, too few to do a reliable calibration. After a few more seconds, even most of those stars disappear. Unfortunately, this means that reference star fields need to be imaged during precious totality time. This procedure reduces the risk in the experiment.
Figure 7. Sky brightness data from two eclipses show that only bright stars are visible outside of the two-minute totality period. Squares are from Möllmann and Vollmer, (2006); dots are from Strickling (2016). Data points are normalized to magnitude 9.5 during totality.

Two reference star fields are required during the brief totality phase: one series taken before the eclipse field, and one taken afterward, on opposite sides of the sun. Immediately after totality begins, the first reference field is imaged 15 times in 30 seconds. The frame is shown on the right side of Figure 8. The figure is shown parallel to the sky’s right ascension axis, and should be rotated counterclockwise by 27° for correct orientation with the horizon. This makes this reference field slightly higher in elevation, but the telescope orientation with respect to the horizon is nearly fixed, minimizing flexure changes. This star field is about 8° west of the sun, in a slightly darker part of the sky. There are 35 stars here brighter than magnitude 10.5, all of which should be measurable. The gravitational deflection averages only 0.06 arcsec, varying by only ± 0.01 arcsec. This particular star field was chosen because the stars fall within a small brightness range; all of the stars should be captured in single exposures.

The first reference field imaging will start just after totality begins. Then, 30 seconds later, the telescope will be repointed to the central eclipse field. Exposures here will be bracketed from 0.2 second to 1 second. After 60 seconds imaging the central eclipse field, the telescope will be repointed to the second reference field for the last 30 seconds of totality, about 8° east of the sun. This field is shifted slightly in declination, again simply to maximize the number of measurable stars. The shift is minor, and is in the direction to minimize the effects of refraction. The field is also not too far from the meridian, to avoid a meridian flip in the MyT Paramount that would take 30 seconds. The time needed to move the telescope from the eclipse field to either reference field, then re-start tracking, is measured at about 3 seconds. Half of the total eclipse time is spent on the reference fields and half of the time on the eclipse field. This reduces the risk in the final analysis.

One additional requirement in the experimental plan is to make sure the telescope is focused as well as the seeing allows. This maximizes the star’s SNR. While focusing on stars during setup on the previous night will give a good starting point, changes in temperature will require a small adjustment, especially since the depth of focus for the Tele Vue NP101is is only about ± 2*λ*(focal ratio)^2, or ± 37 microns for red light. Fortunately, the focuser can easily achieve this resolution using its 10:1 fine focus knob, and locking the focus does not change the distance. The requirement is to find an appropriate star upon which to focus. Based on the curves in Figure 7, magnitude +0.5 Procyon should be visible at least 10 minutes before totality. It is only 36° from the sun, at nearly the same altitude, and will not require a meridian flip. This makes it a perfect target with plenty of time to spare.

3. Star Positions

3.1 Star catalog choices

All previous eclipse experiments required imaging the same star field six months before or after the eclipse, in order to determine the un-deflected star positions to sub-arcsecond accuracy. Parallax was not important, since the geometry of the sun and the earth was the same. This was one of the most challenging parts of their experiments, since the telescope and camera had to be left un-touched for six months in order to minimize mechanical errors. The tests were done at night; the temperature was different, and this also had to be corrected.

Since then, the Hipparcos satellite has provided a very good astrometric reference catalog, apparently obviating the need to measure stars before or after the
eclipse. However, since those measured positions are now 25 years old, the uncertainties for the best stars, the Tycho-2 subset, are typically 0.1 arcsec. More recently, the UCAC4 catalog was released, based on ground-based images taken from 1998 to 2004. Unfortunately, its mean position errors (estimated for 2017) for the stars in the central eclipse field are still about 0.05 arcsec. Neither of these standard catalogs are nearly good enough for this experiment. Fortunately, there are two work-arounds.

The USNO just released their URAT1 catalog last year, with typically 0.01 arcsecond measurement errors. Since only 2.5 years would elapse between the catalog epoch and the 2017 eclipse, this solves most of the problems. Unfortunately, the URAT1 catalog does not include parallax. This makes the position too uncertain for nearby bright stars, unless the right geometry is used. During the eclipse, the stars are in line with the sun, so the parallax is essentially zero. To calibrate telescope optical distortion at night, however, it is best to take images near the zenith, where the atmospheric refraction is minimized. Those stars are not opposite the sun, so some of the stars may have a significant parallax error. The brighter stars, whose measured parallax values can be copied from older catalogs, might be the only allowed stars. In September of 2016, however, the situation becomes dramatically better.

The ESA Gaia satellite is the newest generation of astrometric satellites. It was just launched in 2013, and is in the middle of measuring a billion stars with an accuracy of 0.000024 arcseconds. Since it is still measuring stars, it has not had enough time to measure parallax or proper motion. The ESA has decided, however, to release the first catalog at the end of the summer of 2016! This first catalog will combine the measurements from the Tycho-2 catalog to get very accurate proper motions and parallax numbers. This will be the ultimate solution, just in time for the 2017 eclipse. I am hoping that this schedule can be maintained, and am anxiously awaiting the new catalog.

Figure 9. The Gaia satellite is now measuring stars with micro-arcsecond precision, and ESA will release the first catalog just in time for the 2017 experiment. [ESA]

3.2 Refraction corrections

The precise apparent positions of stars depends on their catalog positions, modified by proper motion, parallax, precession, nutation, stellar aberration, solar gravitational deflection, and local atmospheric refraction [Kaplan 1989]. Rigorous software to combine all of these features was developed by the US Naval Observatory, by the astronomers who co-produce the Astronomical Almanac and the Nautical Almanac. The program is called NOVAS, for Naval Observatory Vector Astrometry Software [Kaplan 2011], and is free to download. It comes in FORTRAN, C, and Python language editions, so one of those compilers needs to run on the user’s computer. NOVAS version 3.1 uses only a simple subroutine to correct for refraction, so I incorporated a more precise routine based on the work of Stone [Stone 1996], also of the USNO (Flagstaff Station). My modified program now outputs stellar positions corrected to a relative refractive error of 0.005 arcsec.

To maintain this precision, I will monitor the local air temperature to within ± 2°F and the atmospheric pressure to within 3 millibars. The FAA weather station nearest the eclipse site can provide the air pressure normalized to sea level, so I will correct it back to the actual pressure at the site’s elevation. I will also measure the air temperature near the telescope with a fast-response thermometer. Since the air temperature falls during the eclipse, an ordinary thermometer won’t work; I designed an electronic thermometer with a 1-second response time in air accurate to 0.2°F, and I will make a recording of the temperature during the entire eclipse. These efforts should make the apparent star positions much better than required.
3.3 Geometric lens distortion

![Graph showing telescope optical distortion](image)

Figure 10. The optical distortion for the Tele Vue refractor amounts to 2 arcsec at the corners of the image. This must be reduced by a factor of 100 to meet the precision requirements of the experiment. A small change in the location of the optical axis makes a big difference in the distortion correction.

All telescopes suffer from lens distortion, since the optical designer prefers to minimize coma, spherical aberration, astigmatism, and field curvature. Fortunately, this geometric distortion can be easily corrected in post-processing. All that is needed is a good measurement of the magnitude of the distortion. However, it turns out that this correction may be the most significant error source in the experiment because locating the precise optical center is difficult.

From Nagler’s optical raytracing, a pretty good estimate of the distortion for the NP101is telescope is possible. I fit the distortion values he provided to a simple quadratic curve, plotted in the next figure, so I could predict the distortion for any star in the field of view. By taking images near the zenith at night, I can compare the measured distortion to the calculated; this is now in process. The coefficients in the polynomial curve are not expected to change over temperature or focus, and these will be verified for this telescope before the eclipse date. Hopefully, once measured, a simple check on the reference fields on eclipse day will be all that is needed to make star position corrections reliably down to the 0.01 arcsecond level.

However, the distortion varies as the square of the distance from the optical axis, and many of the stars are at large distances from the CCD center. Hence, the corrections are very dependent on the precise location of the optical axis on the CCD. In fact, if the optical axis moves only 25 microns from the calibrated position, the error in location will be off by 0.01 arcsec. The important question here is how to either insure that the camera location does not change by more than 25 microns (4.5 pixels on the focal plane), or determine some technique to measure it. Using the reference field data and fitting the star location to the expected (distorted) positions is the simplest method. Since reference fields will be taken on either side of the sun during totality, the location of the optical axis can be averaged to get the location for the eclipse images. This technique, along with others, is currently being tested.

4. Image Analysis

After the eclipse images are ready (dark-frame and flat field corrected), the star locations need to be accurately measured. In past astrometric programs, I used MaximDL [www.maximdl.com] in manual mode. Since there may be only a few hundred stars to measure, this is one option. Each star can be examined to make sure there are no image artifacts, like cosmic rays, that might skew that star’s location. For the calibration data, however, there might be thousands of stars, so some automation is beneficial. In this case, a few bad stars might be ok, since they will be averaged out. The image processing software must be able to measure the star location to 0.01 pixels.

There are several standard methods to determine star locations [Stone 1898; Mighell 1999; Thomas 2004]. The most accurate techniques include simple barycentric calculations and Gaussian curve fitting. The barycentric technique multiplies the intensity of each pixel by its coordinates, and then divides by total intensity. This calculation can be affected by noise, but works for cases where most of the signal is contained in just four pixels. The alternate method fits a 3-D Gaussian curve to the pixel intensities, mathematically looking for the best fit. The location of the Gaussian center is reported as the star location. When only a few pixels are illuminated, this technique is also subject to errors. One improvement would be to constrain the test-Gaussian curve diameter to be the same for every star, but this software is yet to be developed.

Automated software programs, including Pinpoint, Prism, and Astrometrica, were used in the preliminary data analysis. The main problem here is that they use the outdated star catalogs to perform the measurements, leading to small, but important errors. By the time of the 2017 eclipse, I hope to have developed some custom software.
5. Preliminary tests (in Process)

While there is a range of measurements, the consensus is that the sky during totality is about as bright as when the sun is 5.5° below the horizon. The most accurate preliminary test is to image the three star fields during those particular few moments of twilight, and when the star fields are approximately at the same elevation. This occurs in late March and early April. The tests should include the same exposure durations and timing as in the real eclipse. This data will be processed in the next few months using the same software, and the results compared with the best star catalogs. The results should show a gravitational deflection of zero, of course, since the sun is far away. The uncertainty in the measurements is expected to be less than 0.01 arcsec, slightly better than that expected during the eclipse. This proves the technique and gives confidence in the 2017 experiment. Once the Gaia catalog is available, the data may be re-processed. A second dry run will be performed in April 2017, as a final test for eclipse day only four months later.

6. Observation location

The site for the experiment is likely in Wyoming, but a site-selection trip is planned for this August, with the hope that the weather will be similar in 2017. The notorious winds of Wyoming are a concern, not just for the stability of the equipment, but also because the air quality might be affected. The increased transparency for the high altitudes in Wyoming might be degraded by dust stirred up by the wind. Fortunately, modern technology again offers the benefit of pretty accurate 48-hour weather forecasts, so moving the experiment will be easier.

7. Conclusions

This experiment repeats the measurements that made Einstein famous. It is a very difficult experiment because all predictions must come true and no hidden errors must be overlooked. Modern technology, including essentially perfect wide-field telescopes and high speed CCD cameras, along with accurate star catalogs, make this a much simpler experiment than any previous attempts. The anticipated results will also be far more accurate than any previous ground-based attempt. Whether or not this experiment goes as planned, the next USA opportunity will be in Texas in 2024.

8. Acknowledgements

While this experiment was primarily the responsibility of the author, he received critical help and aid from a wide variety of friends and associates. He wishes to thank Al Nagler of Tele Vue for loan of the NP101is telescope and optical raytracing, Greg Terrance of Finger Lakes Instrumentation for loan of the ML8051 camera, George Kaplan and John Bangert for advice on astrometry and URAT1, Suresh Rajgopal for assistance in setting up NOVAS, and Christian Viladrich for his 2006 eclipse FITS image files.

9. References

A Student-Centered Astronomical Research Community of Practice

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Abstract

For over a decade, students from Cuesta College and number of high schools have engaged in astronomical research during one-term seminars. A community of practice—consisting of students, educators, and astronomers—has formed that is centered on supporting the students' astronomical research. The seminar has recently adopted distance education technology and automated telescopes in a hybrid form of on-line and in-person collaborations between students, educators, and astronomers. This hybridization is not only resulting in new areas of growth and opportunity, but has created a number of challenges. For example, as more schools joined this seminar, standardized teaching materials such as a textbook and self-paced, online learning units had to be developed. Automated telescopes devoted to expanding student research opportunities within this community of practice are being brought on line by Concordia University and the Boyce Research Initiatives and Educational Foundation. The Institute for Student Astronomical Research supports this growing community in many ways including maintaining a website and editing books of student papers published through the Collins Foundation Press.

1. Introduction

When it comes to sports, everyone gets it; you have to play to really understand, experience, and learn what the game is all about. It would be ludicrous to teach basketball by practicing basketball fundamentals in the gym (layups, free throws, jump shots, dribbling, and defense), reading about and attending professional basketball games, but never playing in a game. As important as classes and teaching laboratories may be in science education, there is simply no substitute for active engagement in scientific research to show students what science is all about and, perhaps even more importantly, to inspire and motivate them to become scientists or at least appreciate science.

It is a widely held misconception that a student cannot really do meaningful, publishable scientific research until he/she is in graduate school. In actual fact, college undergraduates and even high school students can make original and significant scientific research contributions.

Astronomical research, in particular, is very well suited to engage the beginning high school or college undergraduate researcher. The night sky’s inherent accessibility and also its inherent grandeur are natural draws for the curious student’s mind. And much can be learned and discovered using small telescopes.

In sports, joining a team is a key aspect of the sports experience. Similarly in science, joining a research team and thereby entering a “community of scientific practice” is fundamental and transformational. As important as working with equipment and acquiring data happen to be in scientific research, this is only the beginning of the research process. Student researchers of all ages—particularly high school students and college undergraduates—have much to gain by writing up their results for publication, going through the peer review process, and giving presentations on their research. But this only works if the student researchers are imbedded within the community of practice.

This paper describes a student-centered, small telescope, astronomical research community of practice. This community, centered on the Astronomy Research Seminar offered by Cuesta College (ASTR 299), was described at last year’s
Society for Astronomical Sciences symposium by Johnson et al. (2015), so it will only be briefly summarized in this Introduction.

For eight years, undergraduate student teams have been planning, conducting, and reporting their original research in papers published in the Journal of Double Star Observations and elsewhere. They have presented their results at scientific conferences. These student scientists have conducted research within a supportive, student-friendly community of practice that includes professional astronomers, advanced amateur astronomers, student graduates of the seminar, and experienced educators. Over 40 papers with well over 100 student coauthors have been published within the time constraints of a one-semester community college course or a short summer science camp.

There are, very purposefully, no prerequisites for the seminar. All students are welcome, including high school students signed up as community college students (often taking the seminar as their first college course). We have found that a mix of students actively pursuing technical curricula with students not so involved but interested in astronomy and space works well, and often results in the latter switching their career paths to science. When students join a real-world scientific research community of practice, they become scientists; the transformative experience of being a published scientist can motivate students to consider careers as a scientist. Being a coauthor of a research paper enhances student’s careers through admissions at choice universities, often with scholarships.

The research seminar was, until recently, structured with just a few student teams meeting in person with the seminar’s instructor at Cuesta College’s South Campus. In the spring of 2015, the seminar was restructured as a hybrid in-person/online seminar. Several student teams met with a volunteer assistant instructor at one of 10 separate, distantly located schools in California, Hawaii, Arizona, and Pennsylvania. Instruction was provided through online, self-paced learning units on double star research, planning and managing a scientific research project, writing scientific papers, and presenting research results at conferences. The seminar’s overall instructor (Genet) regularly met online with the individual teams and their volunteer assistant instructors. The entire multi-campus group met together (via Internet meetings) for PowerPoint presentations by each team on their planned research and, near the end of the seminar, their research results.

One team, in San Luis Obispo, California, made measurements of a double star with an astrometric eyepiece (Collins et al. 2016). Another team conducted an educational experiment (Brewer et al. 2016) with astrometric eyepiece observations made by three teams of eighth-grade students at the Vanguard Preparatory School in Apple Valley, California, under the supervision of Sean Gillette. Each team consisted of about a dozen students plus an amateur astronomer or two who supplied the telescope and assisted the students with their observations. A carefully choreographed process and two weekends of hard work resulted in three published student papers (Anderson et al. 2016, Brewer et al. 2016, and Gillette et al. 2016).

![Figure 1: The eighth-grade students at the Vanguard Preparatory School pose with their supporters.](image)

Six teams made double star measurements using “regular” CCD cameras. One team, at Crean Lutheran High School in Irvine, California, made in-person CCD observations with an 11-inch telescope at Robert Buchheim’s Altimira Observatory (Liang et al. 2016), while another team at Leeward Community College in Pearl City, Hawaii, made in-person CCD observations with a 20-inch telescope at the College’s observatory, located just a few feet from the edge of Pearl Harbor (Martin et al. 2016).
Two small teams in Arizona used Richard Harshaw’s 11-inch telescope and high-speed CCD camera at his Brilliant Sky Observatory in Cave Creek to make speckle interferometry measurements of a number of close double stars (Wuthrich and Harshaw 2016, and Dolbear and Harshaw 2016). A team at Lincoln High School in Stockton, California, developed software for plotting orbits of binary stars (Smith et al. 2016a, and Smith et al. 2016b).

Besides publication of the seminar’s papers referenced above in an online journal, the *Journal of Double Star Observations*, the team’s papers were also reprinted as a hardcover book by the Collins Foundation Press (Johnson 2016). While online journals are becoming increasingly popular, it is understandable that students and their supporting institutions still like to see their work in print.

Although the hybrid seminar worked, many areas of the seminar clearly needed further development prior to expanding the scale of the seminar. Based on the lessons learned in the first hybrid seminar, described below, the past year has been spent on seminar support infrastructure as opposed to seminar expansion, preparing for a planned expansion in the fall of 2016. We greatly scaled back the number of seminar teams for the 2015 and spring 2016 semesters to allow us to focus on seminar refinements.
2. **Communities of Practice**

We are developing a somewhat unique educational theory for our seminar that is based primarily on the integration and extension of the social learning theory of “communities of practice” developed by Etienne Wenger (1998) and others. This theory, and its integration into the student-centered research seminar, is discussed in some detail below.

2.1 **Social learning within communities of practice**

“Communities of practice” is both a cornerstone concept in social learning theory and the title of the seminal book in which Etienne Wenger (now Etienne Wenger-Trayner) lays out the theory (1998). A recent definition (Wenger-Trayner, Wenger-Trayner, & de Laat 2015):

Communities of practice are formed by people who engage in a process of collective learning in a shared domain of human endeavor: a tribe learning to survive, a band of artists seeking new forms of expression, a group of engineers working on similar problems, a clique of pupils defining their identity in the school, a network of surgeons exploring novel techniques, a gathering of first-time managers helping each other cope. In a nutshell: Communities of practice are groups of people who share a concern or a passion for something they do and learn how to do it better as they interact regularly.

2.2 **Pedagogy rooted in social learning theory**

Science instruction has traditionally taken a cognitive approach. It has consisted of classroom lectures and laboratory exercises that allow students to move through many subjects within a single discipline, exploring all of the essential knowledge of the field. The students listen and apply concepts in laboratory exercises, often under the supervision of teaching assistants. Laboratory activities develop student skills in measurement, analysis, and report writing. In contrast with this approach—where learning takes place when knowledge is successfully transmitted from a source of knowing to someone who doesn’t know—social learning takes place in the process of becoming a member of a community that defines what competence means in a specific domain of expertise. As Wenger (1998) points out:

Learning is a matter of engagement: it depends on opportunities to contribute actively to the practices of communities that we value and that value us, to integrate their enterprises into our understanding of the world, and to make creative use of their respective repertoires … Practice is a process of interactive learning [that] enables newcomers to insert themselves into existing communities. It is the learning of mature members and of their communities that invites the learning of newcomers (277).

A key pedagogical implication of communities of practice is that they naturally have a spectrum of participants, from core members to more or less peripheral participants. In any specific scientific research community of practice, such as astronomy, there is a gradation of participation and learning that takes place as people move toward full participation. Full participation generally reflects the degree to which the conduct of scientific research is one’s primary vocation. Under the classical model of science instruction, research scientists at research institutes and universities are normally in charge of most research projects. Under their leadership, much of the work on these projects is accomplished by graduate students working on their doctoral degrees (as well as a variety of support staff). Graduate students often also assist research scientists in their teaching duties. More peripheral are serious amateurs, and sometimes undergraduate or even high-school students. Unlike science graduate students working toward their doctoral degree, undergraduate and high school students are not normally granted professional standing and are therefore not given access to the full range of research instrumentation, professional societies, and meetings.

We are endeavoring to change that, and to grant these students a more fully recognized and supported role as peripheral participants in selected scientific research communities of practice. The students in our seminars are learning by developing actual proficiency in the practices of a specific community of scientists, and publishing their work in places where it can be accepted or rejected by experienced members of that community. The process of introducing students to a specific practice of that community (writing a scientific paper) is not simply a matter of teaching them a set of writing conventions that they then apply; it also involves having expert members of that community act as gatekeepers to the dynamic practices of the community by reviewing the students’ papers and providing them critical feedback.

For these students, learning scientific practice is not just learning a skill but developing a new identity, which Wenger (1998) describes as a core dimension of learning in a community of practice:
Learners must be able to invest themselves in communities of practice in the process of approaching a subject matter. Unlike in a classroom, where everyone is learning the same thing, participants in a community of practice contribute in a variety of interdependent ways that become material for building an identity. What they learn is what allows them to contribute to the enterprise of the community and to engage with others around that enterprise (271) … Learning [within a community of practice] transforms our identities: it transforms our ability to participate in the world by changing all at once who we are, our practices, and our communities (227).

More recent work in social learning theory has provided two additional resources for our work. On a practical level, recent work has produced a formal framework to evaluate the type of learning experienced in our seminars, in which the knowledge gained is based on a mixture of qualitative and quantitative data (Wenger, Trayner, and de Laat 2011). On the theoretical level recent work has situated learning in communities of practice within a broader landscape of practice (Wenger-Trayner and Wenger-Trayner 2014).

In a complex, multidisciplinary world, the boundaries between scientific disciplines also become potential places of learning, as in multi-disciplinary approaches to problem solving and exploration, or in new scientific areas where traditional disciplines begin to overlap and sometimes converge. In this case, a student’s developing a sense of the boundaries of a community of practice is a significant part of the learning process. This is all the more important because some students may never become scientists in their core discipline, but rather, they may develop an identity that traverses several communities in that landscape.

As Wenger (1998), suggests, the benefit of involving peripheral participants also serves the community. Under the classic learning model:

… there are all sorts of reasons to shelter newcomers from the intensity of actual practice, from the power struggles of full participation … Similarly, there are all sorts of reasons to shelter old-timers from the naivété of newcomers and spare them the time and trouble of going over the basics. … When old-timers and newcomers are engaged in separate practices, they lose the benefit of their interaction. … Communities are thus deprived of the contributions of potentially the most dynamic, albeit inexperienced, segment of their membership — the segment that has the greatest stake in their future.

2.3 Developing and spreading a new pedagogical practice

Communities of practice constitute both a learning theory and a practical approach to enabling learning (Wenger et al. 2002). By involving educators in the development of our seminar, we are creating the kernel of a community of practice of educators and researchers who learn together how to make this approach to the teaching of science work. This means that beyond our seminar, this community of educators and researchers has the potential to continue the development of this pedagogical practice and spread the approach by inviting others to join their community (Wenger, McDermott, & Snyder 2002).

It could be beneficial if more students joined, early-on, the communities of practice most relevant to them. As Wenger (1998) suggests:

… schools gain relevance not just by the content of their teaching—much of which can be acquired just as well in other circumstances—but by the experiments of identity that students can engage in while there. Consequently, deep transformative experiences that involve new dimensions of identification …even in one specific or narrowly defined domain – are likely to be more widely significant in terms of the long-term ramifications of learning than extensive coverage of a broad, but abstractly general curriculum. … This is especially true in a world where it is clearly impossible to know all there is to know, but where identity involves choosing what to know and becoming a person for whom such knowledge is meaningful.

3. Lessons Learned and Seminar Limitations

The community of practice approach that we have evolved enables us to meet the standards of quality research and replicability embraced by the classic graduate research science-student learning model while, at the same time, transforming the identities of undergraduate and even high school students long before they reach graduate school. We have found, over the years, that the success of student research projects results primarily from simply following the rules of scientific research. Under the classic learning model these practical rules are often only learned as a graduate student, post-doctoral researcher, or even later. Frustrated that so many doctoral students were graduating without learning the basic “ropes” of being a scientist, Peter Feibelman (2011) developed a course he turned into the classic book, A PhD is not enough! A Guide to
Survival in Science. Although we initially tried “softening” science’s rules for seminar students, we found this undermined the core strength of the research seminar, so we returned to the full application of science’s rules to our seminar. Science, a highly successful mode of cultural evolution, has developed its many rules for good reasons. Students like to know that they are doing the real thing, not some watered-down version that they are likely to interpret as condescending.

3.1 Follow the regular “rules” of classic scientific research

- Research is conducted within a community of practice, typically within a narrow frontier specialty.
- Research must be original and be published as papers in appropriate (specialty) journals that are reviewed and read by the members of the relevant community of practice.
- Publication is mandatory and places everyone’s reputation on the line, including that of the students, instructors, schools, journal, and the seminar itself. Thus papers need to be of high quality.
- Publication is not enough. Results should, if possible, be presented at conferences attended by researchers from the relevant community of practice.
- Team members are not expected to contribute equally. Author order provides justice to variations and allows each member to contribute as their time, talents, knowledge, and experience dictate.
- Team members should be included or added, as needed or desired, from outside the seminar.

In honing our approach into a framework with demonstrated success, we have learned some valuable lessons.

3.2 Practical lessons learned over the years from our community of practice approach

- Students require most of the last half of the term to write and rewrite their papers. Thus each team needs to plan on completing their observations, data reduction, and analysis in the first half of the seminar.
- Student teams must not be allowed to fail, as this would impede their careers instead of advancing them, and would give the seminar a very negative reputation among students. The most frequent cause of failure is taking on too much. Instructors must help the students limit the scope of their projects such that they are achievable within the time constraints of an academic term.
- Instructors should help teams avoid “mission creep,” i.e. adding additional research objectives.
- It is sometimes appropriate for the instructor to bring in outside helping experts, add team members with special skills, or call in an experienced “paper rescue team” to mentor (or tutor) the students through the final stages of their project. If the outside rescue team contributes significantly to the project (which is likely), they should, of course, be included as paper coauthors.

While the community of practice framework has had demonstrated success, its applicability and usefulness so far has some important limitation around which we are focusing some of our current research to identify potential workarounds.

3.3 Seminar limitations

- The hybrid version of the research seminar has only been taught by a very experienced instructor assisted by hand-selected, already knowledgeable, and often highly experienced double star researchers. It is not clear that new, and somewhat naive instructors or assistant instructors, could be efficiently trained in, for instance, summer workshops. An ability to economically train instructors and provide them with appropriate material for their classes is a prerequisite for more widespread propagation of the seminar.
- Only one specialty area of research within astronomy (double star astrometry) has been significantly pursued. This area was chosen because it can be done with low cost of instrumentation, the measurements are fairly straightforward, it is conceptually simple, and (perhaps of greatest import) a research project can be completed, including a paper submitted for publication, within the demanding confines of a single academic term or summer science camp. There is, for this specialty, a dedicated forum for publishing results with a relatively rapid
turnover, the *Journal of Double Star Observations*.

- Although the seminar has been taught for many years, the only basis for its evaluation, so far, has been its track record of published papers, conferences, and books—a public record accessible to all for evaluation. Although the instructor and the volunteer assistant instructors have their opinions on what factors have influenced student outcomes in favorable (or unfavorable) ways, no formal assessments—even simple ones—have ever been made.

4. The Institute for Student Astronomical Research

The Institute for Student Astronomical Research (InStAR) was chartered in 2015 as a division of the Collins Educational Foundation, a 501(c)(3) non-profit organization. In addition to the support provided by its parent organization, InStAR has received extensive support from the Boyce Research Initiatives and Educational Foundation (BRIEF) as well as a number of other organizations and individuals. The primary objective of InStAR is to help organize and support a student-centered small telescope astronomical research community of practice. Emphasis is placed on facilitating and supporting published astronomical research by high school and undergraduate students.

Specific InStAR objectives are to: (1) develop training and assessment material, (2) maintain the InStAR website, (3) develop and maintain a central computational server, (4) organize conferences and workshops, (5) support student observations on robotic and large mountaintop telescopes, (6) publish books in support of student research, and (7) obtain financial resources to support student research.

Each of these InStAR activities is detailed below.

4.1 Develop Training and Assessment Material


** Expand this book to include other areas in astronomy beyond double stars that are suitable for small telescopes and have a solid pro-am support community. These include eclipsing binary, variable star, asteroid, and exoplanet transit photometry, asteroid and lunar occultations, and carefully selected areas within spectroscopy such as low-resolution spectral classification.

** Create and refine other training material including guides for using robotic telescopes, specific software suites, and specific calibration procedures.

** Develop instructor training material.

** Develop material to assess the effectiveness of teaching students the basics of scientific research and evaluate the ability of the seminar to change student attitudes toward science and pursuing science, mathematics, or engineering as a career.

4.2 Maintain the InStAR Website

** Develop, refine, and maintain the InStAR website (this is kindly supported by the Boyce Research Foundation) to provide support material for student research and be the repository of videos, papers, and software.

** Complete the design of the website to have three compartmented divisions with different restrictions on access: student resources, instructors and facilitators resources, and unrestricted public resources.

4.3 Develop and Maintain a Central Computational Server

** Provide online software tools for analysis.

** Provide a repository for observational data.

4.4 Help Organize Conferences and Workshops

** Initiate and develop full conferences, such as the Concordia International Conference on Close Binary Stars, July 2017.

** Initiate and develop meetings in conjunction with others. An example is AAS “meeting within a meeting” on High School and Undergraduate Research in Astronomy, June 2016.

** Initiate and organize workshops and conferences where students present research results, teachers are trained, and the continued development of InStAR function is a topic. A number of workshops have already been held, the most recent being the Concordia Workshop on Student Astronomical Research and Robotic Telescopes in June 2015.
4.5 Support Student Observations on Robotic and Large Mountaintop Telescopes

** Purchase blocks of time on robotic telescope networks for student observations, including both CCD and speckle interferometry technologies. Supply (on extended loan) an EMCCD camera for making the speckle observations.

** Arrange student observational opportunities on larger telescopes. Submit proposals to public observatories. Purchase blocks of time for observations where appropriate. Arrange the observational sessions.

4.6 Publish Books in Support of Student Research

** Support the writing, editing, and publishing of astronomy research-related books by the Collins Foundation Press, another division of the Collins Educational Foundation. These low cost books will be provided in both paper and e-book versions.

** Included will be books containing student papers, some of which will be reprints of special issues of the online Journal of Double Star Observations.

** Also included will be books with material from conferences and workshops.

Relevant books currently available from the Collins Foundation Press:


*Speckle Interferometry of Close Double Stars,* Eds. R. Genet, E. Weise, and V. Wallen, Foreword by David Rowe.

*Double Star Research: A Student Centered Community of Practice,* J. Johnson, Foreword by John Kenney.

*Double Star Research and Development Summer Seminar,* J. Haas.

4.7 Obtain Financial Resources to Support Student Research

** Obtain and budget additional funds, including donations from individuals and funding by foundations and governmental organizations. InStAR’s parent, the Collins Educational Foundation, manages this function.

** Prepare and submit proposals for grants.

5. Seminar Expansion and Alternative Venues

With a textbook, self-paced learning units, a web site, a software server, robotic telescopes coming on line, and a supporting organization (the Institute for Student Astronomical Research), we have felt that we can not only continue the seminar, but expand it. We also think it would be useful to experiment with alternative venues.

For four years in a row (2008-2011), the seminar was conducted as an informal (non-college credit) student research seminar in conjunction with the University of Oregon at their Pine Mountain Observatory (PMO) east of Bend, Oregon. Students from high schools and universities across the Pacific Northwest participated in this informal summer seminar. Their research was primarily double star astrometry, though a few teams led by Richard Berry did photometry of variable stars. The students’ research resulted in a number of published papers.
This summer (2016), we will be offering a six-week hybrid (online and in-person) seminar through Cuesta College for credit. Seminar student participants are being solicited from Cuesta College itself, nearby California Polytechnic State University (Cal Poly), and Apple Valley High School. The Cuesta and Cal Poly students will work together in teams which will feature team members from both institutions. They will be making observations of double stars in the southern hemisphere using robotic telescopes in Australia (iTelescope). The students at Apple Valley High School (including some of the eighth grade students who are already published double star researchers and are now high school students) will be using astrometric eyepieces and, weather permitting, will be making their observations at the fourth annual Apple Valley Double Star Workshop under the leadership of Mark Brewer.

The San Diego area is being targeted for a major expansion of the seminar starting this coming fall (2016). The cadets at the Army and Navy Academy, under the direction of Pat and Grady Boyce, have had the most experience with the new hybrid online/in-person seminar format. The Boyce Research Initiatives and Educational Foundation (BRIEF) is fostering the expansion of the seminar in the San Diego area. This fall (2016) there will be student teams from several high schools in the San Diego area participating in the hybrid Astronomy Research Seminar.

6. Pushing Double Star Astrometry to the Limit

For many years, Cuesta College’s Astronomy Research Seminar featured visual (eyeballs at the telescopes) astrometric eyepieces as its primary instrument. Observations were made of relatively bright and wide double stars using telescopes provided and operated by local amateur astronomers. The students made and recorded the astrometric observations. They enjoyed making telescopic observations with the assistance of amateur astronomers. The instrumentation was simple, low cost, and rugged.

Astrometric eyepiece observations of relatively wide and bright double stars have continued as an observational mode in the seminar, especially by students in the Apple Valley area. The use of video recordings of the observations has added a level of sophistication and precision to these intuitively pleasing observations. Visual observations have also been made with a double-image Lyot micrometer (Weise et al. 2015).

It was only natural, however, that some student research teams would try other double star observational modes. Observations with CCD cameras allowed much fainter doubles to be observed with higher precision. This opened up, for instance, a large number of faint doubles for observation. While CCD instrumentation and reduction was more complex, these observations were, nevertheless, appealing to many teams.
Since most automated telescopes are equipped with CCD cameras, they can be used to make the
observations, overcoming problems with local weather, ownership of equipment, and, quite often,
initial reduction with respect to darks and flats. The cadets at the Army and Navy Academy have used the
automated telescopes at iTelescope to good effect on a number of their team projects.

Observation of double stars with separations below the seeing limit via speckle interferometry have required, until very recently, the use of high speed, low readout noise, electron-multiplying CCD cameras. These emCCD cameras are expensive ($14,000 and up) but, thanks to grants from the American Astronomical Society and California Polytechnic State University, we were able to acquire two Andor Luca emCCD cameras.

Their use required magnification (Barlow lenses) which results in very small fields-of-view. In spite of their expense and very small fields-of-view, we were able to observe doubles with separations of only about 0.5ʺ on smaller telescopes, such as the 20-inch PlaneWave Instruments Corrected Dall-Kirkham at Pinto Valley Observatory (Genet et al. 2015a), and 0.1ʺ on the 2.1-meter telescope at Kitt Peak National Observatory. While observations of such close double stars opened up many scientifically exciting possibilities, observations were limited by equipment expense, observational difficulties, and (for the closest doubles) the availability of larger telescopes for student observations.

In a very recent technical breakthrough by Sony, high speed, low noise CMOS chips have become available to camera manufacturers. ZWO was the first company, with its ASI 224, to place one of these Sony chips into a camera (Genet et al. 2016 and Ashcraft 2016). A similar and even better performing Sony monochrome chip has just been made available by ZWO as their ASI 290 camera. These ZWO cameras cost about $400, and are much smaller and are more rugged than emCCD cameras. Extensive tests of these new cameras—especially the ASI 290—has established that for astrometry they perform just as well or better than the expensive emCCD cameras.

Not only has there been a major breakthrough in cameras for speckle interferometry, but there have been major advances in the availability and convenience of speckle interferometry reduction software. Dave Rowe extended his Plate Solve 3 (PS3) software to include speckle interferometry reduction of close double star observations. This software features low and high pass filters, an interference (axis) filter, and accommodation of single deconvolution stars to remove many atmospheric and telescopic aberrations (Rowe and Genet 2015). The software can also be run in manual and semiautomatic modes; the latter is very convenient for reducing large-scale runs.

Recently, the speckle interferometry portion of PS3 has been issued as its own, independent software, the Speckle Tool Box (STB). STB incorporates many refinements and, in a major addition (currently under beta test) is incorporating bispectrum analysis which should allow differential photometry of very close double stars. Multi-band differential photometry of close double stars will facilitate a whole new level of astrophysical analysis for very close double stars.
While speckle interferometry overcomes seeing limitations, the ultimate limit on how close double stars can be successfully measured is determined by the aperture of the telescope—the larger the telescope, the closer the double stars resolvable by speckle interferometry. Our two week-long runs on the 2.1-meter telescope at Kitt Peak National Observatory were very productive, and the mountaintop environment was highly inspiring to the students from a number of schools. However, obtaining observing time was a long, difficult, and highly competitive process. Furthermore, Kitt Peak was quite some distance from California, and cost for room and board was $90/night. This limited student participation, particularly for high school students. Furthermore, the 2.1-meter telescope was mothballed shortly after our second run, and is no longer available for our use (it has reopened as a Caltech adaptive optics telescope using Robo AO).

For our larger telescope, mountaintop observations, we are currently planning to use the 30-inch telescope at Stony Ridge Observatory (SRO), which is conveniently located in southern California not far from Mt. Wilson Observatory (MWO). As with MWO, the seeing at SRO is often excellent, and the mountaintop view of the greater Los Angeles area is quite spectacular. Several of our group’s members have joined SRO as Associate Members.

We are also planning on making limited use of the nearby 100-inch telescope at Mt. Wilson Observatory. Although there is a significant charge for using this telescope, and a modest charge for dorm rooms and the use of kitchen and other facilities, we hope to obtain funding for occasional runs on the famous 100-inch Hooker telescope. We have, already, had an engineering checkout, successfully making speckle interferometry observations on the 100-inch (Genet et al. 2016). Selected (star) students from various research teams would be invited to participate in the 100-inch runs as a partial reward for their hard work.
Finally, we have, for some time, been investigating ways to observe close double stars with faint, late-M secondaries. Late-M stars are the most numerous stars in the universe, but their masses are not well determined because the astrometric observations required to define their orbits, and hence their masses, are difficult. We have, with limited success, investigated shaped aperture masks to direct the light from the bright primary star away from discovery zones to reveal the faint, close secondary stars (Foley et al. 2015).

We are currently evaluating near infrared cameras made by Raptor. While expensive—over $20,000—these cameras are comparatively low cost, small, and relatively easy to use (thermoelectric cooling) as compared to large, very expensive, NIR cameras with LN$_2$ cooling. Jocelyn Serot, in France, has begun the evaluation of one of these cameras, as has Richard Harshaw in Arizona. Many late-M secondaries should appear reasonably bright in the J and H bands with these high speed near-IR cameras.

7. **Moving Beyond Double Star Astrometry with Automated Telescopes**

There are, of course, many areas of astronomical research besides double stars that are suitable for research with smaller telescopes. Early in the seminar we conducted research on intrinsically variable stars with in-person student observations. Photometric observations of several relatively short-period variable stars were made over a number of weeks. By the time the data was reduced and analyzed the semester was nearly over, pushing paper writing beyond the end of the course. Having many nights of observing required lots of driving time by the students’ parents, not to mention time spent by the instructor helping the students with their observations, protecting complex equipment from novice students, etc.

An exoplanet transit was also observed in 2008 by two students of the seminar aided by exoplanet hunter Cindy Foote and a local amateur astronomer, Tom Smith. Exoplanet transits are an excellent candidate for student projects because they usually can be observed from ingress to egress in a single night. Although exoplanet transits are extremely interesting to students, in large-part due to their perceived relationship to habitability and astrobiology, obtaining observations had the same challenges as other photometry projects.

We found that astrometric observations of a double star with an astrometric eyepiece could be made in a single evening. The evening had to be clear, moonlit nights did not work well with visual observations, the students still had to be transported to the observatory or the telescope brought to the school, and there could be equipment difficulties. In spite of these difficulties, we still find this to be an attractive way for students to gather data at locations with good weather, reasonably dark skies, and supporting amateurs with telescopes.

As discussed above, observations have been increasingly made by way of CCD cameras on automated telescopes. This has been mainly due to
increased observational precision and the ability to observe much fainter and often more scientifically interesting double stars. Using automated telescopes obviated the need for student transportation. Since automated telescopes are generally located at good-weather sites, and networks such as iTelescope have multiple telescopes, long waits for good weather are generally not a problem with these telescopes. The use of automated telescopes can obviate the need for schools to maintain and operate telescopes themselves. Furthermore, initial data reduction, such as taking care of darks and flats, are often taken care of by the automated telescope systems.

Using commercial robotic telescopes, such as iTelescope, is usually affordable for astrometry, since just a few, relatively short-exposure images are required. On the other hand, time series photometry of eclipsing binaries, intrinsically variable stars, exoplanet transits, or asteroids can take many hours, not only making the cost prohibitive, but making it difficult to schedule such a large block of time.

One solution to this difficulty (albeit neither easy nor low in cost) is the installation and maintenance of automated telescopes devoted almost exclusively to student astronomical research. The Astronomical Research Seminar is fortunate that four such telescopes will be coming on line shortly to fill this need. Two are at Concordia University, located in Irvine, California. They were funded by a grant from the Keck Foundation that stipulates that much of their use would be devoted to student researchers. One of these telescopes is a Celestron C-14 optical tube assembly (OTA) on a Paramount German equatorial mount, while the other is a PlaneWave Instruments CDK-17 OTA, also on a Paramount. Both are under domes on the Concordia University campus.

The other two automated telescopes, funded by the Boyce Research Initiatives and Educational Foundation (BRIEF), will be located under a roll-off roof in the mountains east of San Diego. One of the telescopes is a Celestron C-11 OTA, donated by the Collins Educational Foundation, on Paramount German equatorial mount (GEM), while the other is a Hubble Optics CDK-20 OTA on a Mathis equatorial fork mount.

Besides being usable with “regular” CCD cameras for time series photometry, several of these automatic telescopes will be equipped with instrument selectors. This will allow switching to high speed cameras located behind Barlow lenses and filters for speckle interferometry astrometry and multi-band photometry (with bispectrum analysis) of close double stars. Instrument selectors should also, in the future, allow switching to spectrographs ranging in resolution from simple grism spectrographs to much higher-resolution fiber-fed spectrographs.

8. Conclusions

A student-centered, small-telescope research community of practice has been developed over the years to support Cuesta College’s Astronomical
Research Seminar. Recently, the seminar has adopted distance learning technology to facilitate the spread of the seminar to many more, geographically dispersed schools. The addition of several automated telescopes dedicated to student research will facilitate expansion beyond double star astrometry to time-series photometry of eclipsing binaries, exoplanets, intrinsically variable stars, and asteroids.

9. Acknowledgements

We are grateful to Cuesta College, California Polytechnic State University, and Concordia University for their support of the Astronomy Research Seminar. We are also grateful to the Collins Educational Foundation, the Boyce Research Initiatives and Educational Foundation, the W. M. Keck Foundation, and Gravic Labs for their generous support of the seminar. Finally, we thank Kitt Peak National Observatory, Pinto Valley Observatory, Altimira Observatory, and iTelescope for making their facilities available for student observations.

10. References


Empirical Measurements of Filtered Light Emitting Diode (FLED) Replacements

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Abstract:
Low pressure sodium (LPS) public lighting, long favored by astronomers and dark sky advocates, is in decline due to a variety of economic issues. Light emitting diode (LED) technology is a rapidly ascendant mode of lighting in everything from residential to commercial applications. The resulting transition from LPS to LED has been accompanied by great angst in the environmental community, but very little has been done in the way of empirical measurement of LEDs in the field and their actual impacts on communities. The community of Waikoloa Village, Hawaii is located on the western slopes of Mauna Kea, within direct line of sight view of the major astronomical observatories on the mountain summit. Waikoloa has been rigorously illuminated almost exclusively by LPS for many years in acknowledgement of the importance of the Mauna Kea Observatories to the Big Island of Hawaii. As LPS ceases to be a viable alternative for local government support, a decision has been made to experimentally retrofit all of the Waikoloa street lighting with filtered light emitting diode (FLED) fixtures. This action has rendered Waikoloa Village a unique laboratory for evaluating the effects of such a change. STEM Laboratory has been awarded a research grant to make a variety of measurements of the light at night environment of Waikoloa Village both before and after the street light retrofit program. Measurements were conducted using a combination of techniques: Satellite Data Surveys (SDS), Ground Static Surveys (GSS photometry), Ground Mobile Surveys (GMS photometry), Airborne Surveys (ABS photography), and Spectroscopic Surveys (SpecS). The impact of the changes in lighting sources was profound, and the preliminary results of this extensive program are discussed in this presentation.
Abstract

Contact binary star systems have been studied for well over one hundred years. Visual and photometric data have been collected and orbital periods have been calculated on hundreds of systems. One would like to know if those observed periods are stable or change over a long period of time. Changes in the period imply an ongoing evolutionary process in the binary system. One method which yields insight into this evolutionary process is by graphing observed data (O) – calculated data (C). These diagrams require period data collected over a considerably long time. Data mining the internet for these O-C diagrams is an effective method of collecting this long term information quickly. Possible evolutionary aspects of the binary system can then be explored; using O-C diagram analysis obtained in this manner.

1. Introduction

We take it for granted that W Ursae Majoris contact binary star systems are actively evolving. To determine if and when changes are occurring, period data over many epochs is required. Observed data (O) – calculated data (C) diagrams yield information on these long term period changes in these systems. We can utilize these diagrams to explore the possible evolutionary processes that are ongoing.

The problem arises due to the fact that to construct an O-C diagram one needs period data over many epochs, possibly up to one hundred years. This data is available, however, through multiple sources on the internet. Data mining allows quick retrieval of this information. Analyzing these diagrams for several star systems, evolutionary changes can then be determined and investigated.

For this study, 62 W Ursae Majoris (W UMa) type contact binary system O-C diagrams were data mined and compared to see if any possible evolutionary processes are ongoing in the systems.

2. The O-C Diagram

If a binary star system is perfectly periodic, every period is the same, we therefore can accurately predict future occurrences. If the system reaches a minimum brightness at time $t_0$ (the epoch), and has a period $P$, then we know the next minimum will occur at time $t_0 + P$. The next minimum after that will be at time $t_0 + 2P$ and so forth. In fact, if we choose $t_0$, as our epoch, $t_0$ will be the time of minimum for cycle number zero, and the computed time of minimum for any cycle number $n$, which we can call $C_n$, is easy to calculate:

$$C_n = t_0 + nP$$

With this formula, we can therefore compute the times of all minima, past, present, and future.

However, these times will only be correct if the system is perfectly periodic. In addition to the computed times of minimum $C_n$, we can also directly observe the stars to estimate the observed time of minimum for cycle number $n$, which we will call $O_n$.

We can now compare theory (the computed times $C_n$) to our observations (the observed times $O_n$). This is accomplished by simply taking the difference between the observed and computed times of minima. These are the “O–C,” or “observed minus computed” values. For each cycle number $n$, we have:

$$(O-C)_n = O_n - C_n$$

After we have determined the O–C values, we can plot O–C as a function of cycle number $n$. This gives us an O–C diagram; a powerful tool for period analysis.\[^1\]
3. O-C Diagram Interpretation

The behavior of binary systems can be interpreted by examining the O–C diagrams themselves. [1] Let’s look at some hypothetical data for several contact binary systems to obtain insight into any observed changes.

In System A (Fig 1) all the O–C values are zero, this is because the theory matches our observations and we have perfect periodicity.

In System B (Fig 2) the O–C values lie on a straight line which is horizontal (parallel to the x-axis), but are all displaced from 0 by the same amount, the system is periodic, and our period is correct, but the epoch, $t_0$, is wrong.

In System C (Fig 3) the O–C values lie on a straight line; but the line is not horizontal, the system is periodic but our estimated period is not correct. When the system is periodic but our period is incorrect, the slope of the line through the O–C values is the difference between the true and estimated periods. In addition, the intercept of the line is the difference between the true and estimated epochs.

In System D (Fig 4) the O–C values leave one straight line and start another with the same slope, but are offset. The period has remained the same but the epoch has shifted.

In System E (Fig 5) the O–C values change from one straight line to another which has a different slope, the period has changed. The slope of each line is the difference between its period and the estimated period.

In System F (Fig 6) the O–C values do not follow a straight line; the system is not strictly periodic, but the period is changing in a regular way.
Things are not always so simple for real data. Many systems are not perfectly periodic. For Mira type variables, for example, the period of each cycle is a little different, although the average period is stable.

Also remember that for all real data, there are observational errors, no matter how precise the instrumentation. Standard error analysis needs to be implemented.

4. W Ursae Majoris Contact Binaries

In 1903 Gustav Müller and Paul Kempf discovered that the star W Ursae Majoris was variable and has since become a prototype of this class of variable stars.

The class is divided into four subclasses: A-type, W-type, B-type and H-type:

A-type systems are composed of two stars both hotter than the Sun, having spectral types A to F, and periods of 0.4 to 0.8 day. The difference between the surface temperatures of the components is less than several hundred kelvin.

W-types have cooler spectral types of G to K and shorter periods of 0.22 to 0.4 day.

B-type systems, introduced in 1978, have larger surface temperature differences.

H-type systems, discovered in 2004 by Sz. Csizmadia and P. Klagyivik, have a higher mass ratio and have extra angular momentum.

Large star spots (bright/dark) have also been observed on the surface of these stars. These spots have been observed through the analysis of multi-color photometric light curves of the systems.

W UMa light curves differ from those of classical eclipsing binaries, undergoing a constant ellipsoidal variation rather than discrete eclipses. This is because the stars are gravitationally distorted by one another (see Fig 7), and thus the projected area of the stars is constantly changing.

The W UMa variable class, also known as low mass contact binaries, are close binaries that share a common envelope of material. They transfer mass and energy through the connecting neck, although R.E. Wilson and others argue that the term "overcontact" is more appropriate.

We can define the Roche lobe as a region around a star in a binary system within which any orbiting material is gravitationally bound to that star (see Fig 8). When a star "exceeds its Roche lobe", its surface extends out beyond its Roche lobe and the material which lies outside that point can "fall off" into the other object's Roche lobe via the first Lagrangian point. In binary system evolution this is referred to as mass transfer via Roche-lobe overflow.

Figure 7: W UMa Contact Binary
W UMa itself consists of two stars plus a distant 12th magnitude companion star (ADS 7494B). The group appears to be moving together through space. The primary component has a larger mass and radius than the secondary, 1.19 solar masses and 1.08 solar radii. The secondary has 0.57 solar masses and 0.78 solar radii. A light curve for W UMa (see Fig 9) shows the characteristics unique for this prototype class of contact binaries. The depths of the brightness minima are usually almost equal since both stars have nearly equal surface temperatures of about 6200 K. The stars have a circular orbit with a period of 0.3336 days, or eight hours and 23 seconds. The maximum magnitude of the pair is 7.75. During the primary eclipse, the net magnitude drops by 0.73, while the secondary eclipse causes a magnitude decrease of 0.68. The author has obtained period and minima values on W UMa over many years which yields consistency with these values but show slightly longer periods, by a few seconds.

Because W UMa have their outer envelopes in direct contact, they therefore have the same stellar spectral classification of F8Vp. The contact nature of these stars makes it impossible to precisely tell when an eclipse of the primary by the secondary starts or ends.

5. W UMa System Evolution

A satisfactory theory for the origin, structure, and evolution of W UMa variables is not yet complete. Li et al. have argued that most W UMa systems formed from detached binaries; taking about 3.23 billion years to progress from detached stars to a contact binary system. It is theorized that the stars in the systems will eventually merge into a single rapidly rotating star. While in the overcontact phase, evolutionary processes are ongoing. The orbital periods change over time as these processes continue. The evolutionary status and dynamical evolution of W UMa systems have been investigated. It appears that there is no evolutionary difference between the W and A type subsystems. Mass - radius (M-R) and mass – luminosity (M-L) diagrams of W and A type systems indicate that a large amount of energy should be transferred from the more massive to the less massive component, so that they are not in thermal equilibrium and undergo thermal relaxation oscillation. Moreover, the distribution of angular momentum, together with the distribution of the mass ratio, suggests that the mass ratio of the observed W UMa systems decreases with decreasing total mass. This could be the result of the dynamical evolution of W UMa systems, which suffer angular momentum loss and mass loss as a result of the magnetic stellar wind. Consequently, the tidal instability forces these
systems towards lower q values (a measure of internal friction) and finally to rapidly rotating single stars.

The orbital period is very sensitive to any effect and can be presently measured with very high precision down to $10^{-9}$ part of the period. A large number of series of minimum observations exist for many system, some going back to the 19th century. [18]

Looking at an O-C diagram for W UMa (see Fig 10) we see obvious changes have occurred in the periodicity over many epochs.

![Figure 10: O-C Diagram for W Ursae Majoris][3]

It appears that from 1903 until about 1920 and again from about 1982 to the present the period was stable but was incorrectly recorded. From 1920 until 1982 dramatic changes were occurring which must be the result of evolutionary changes within the system.

6. Data Mining

O-C diagrams for 62 W UMa type systems were data mined. 33 W type and 29 A type (see Table 1) systems where chosen for comparison to see if these systems exhibit period changes (are evolving) in a similar manner to W UMa. [3] The O-C Diagrams are presented in Table 2 (at the end of the paper) for reference.

<table>
<thead>
<tr>
<th>W Ursae Majoris Systems</th>
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<tbody>
<tr>
<td>AB And</td>
</tr>
<tr>
<td>AC Boo</td>
</tr>
<tr>
<td>VW Cep</td>
</tr>
<tr>
<td>CC Com</td>
</tr>
<tr>
<td>RW Dor</td>
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</tr>
<tr>
<td>ER Ori</td>
</tr>
<tr>
<td>AE Phe</td>
</tr>
<tr>
<td>AH Vir</td>
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</table>

<table>
<thead>
<tr>
<th>Type A</th>
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<tbody>
<tr>
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<td>AH Aur</td>
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<tr>
<td>RR Cen</td>
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<tr>
<td>V401 Cyg</td>
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<tr>
<td>AP Leo</td>
</tr>
<tr>
<td>V508 Oph</td>
</tr>
<tr>
<td>AU Ser</td>
</tr>
<tr>
<td>AG Vir</td>
</tr>
</tbody>
</table>

7. Results

In reviewing the O-C diagrams of the 62 W UMa collected systems; 15 systems (see Table 3) appear to be quiescent and have stable periods during the duration of the data collected. In a few cases there are minimal data points or gaps in the data. The remaining systems all appear to be undergoing period changes similar to W UMa as expected.

<table>
<thead>
<tr>
<th>Stable Period Systems</th>
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</thead>
<tbody>
<tr>
<td>GZ And</td>
</tr>
<tr>
<td>LS Del</td>
</tr>
<tr>
<td>V781 Tau</td>
</tr>
<tr>
<td>AH Aur</td>
</tr>
</tbody>
</table>

One system in particular, XY Leo (See Fig 11), was found to be undergoing radical changes over very short time intervals during the last 80 years. XY Leo is a member of a visually unresolved quadruple system composed of a contact binary and a detached, non-eclipsing, active binary. [21] The gravitational
interaction between these other bodies results in this unusual O-C Diagram.

Figure 11: O-C Diagram for XY Leo [3]

A literature search was also conducted to see what explanations have been presented to illustrate these period anomalies. As explained earlier it was discovered that changes in the period could be the result of several possible ongoing processes:

a) Apsidal rotation (advance of the periastron). [22]
b) Possible third body, stars or planets (up to 50% of W UMa stars have companions). [16] This appears to be the cause of the rapid variations in the XY Leo system. [21]
c) Mass transfer and ejections (angular momentum redistribution changes the period). [14][15][19]d) Cyclical magnetic modulation of the primary star with continuous mass transfer. [18]
e) Braking effects of the magnetic fields (strong chromospheric and coronal X-ray emissions have been detected, indicating a strong, dynamo-related magnetic activity for the system. This magnetic activity is common to W UMa variables). [14]
f) Oscillations in the rotation of close binary stars. [23]
g) Differential rotation of cool stars. [24]
h) Gravitational waves. [25]

8. Conclusions

It is apparent from a review of the data, that W UMa overcontact binary systems are evolving. Their evolution seems to be ongoing on a rapid time scale although several systems have been caught in a relatively quiescent phase. The evidence seems to corroborate the suggestion that these overcontact systems are evolving toward a single rapidly rotating star.

From the literature search, the cause of the observed period changes appear to be predominately from the transfer of mass from the larger star to the smaller one resulting in the redistribution of the angular momentum. Mass ejection from magnetic effects also play a lesser role in the period changes.

Several other phenomenon have also been suggested for the period changes; differential rotation, asteroseimology and gravitational waves. Further research is needed to develop a satisfactory theory for the origin, structure, and evolution of W UMa variables.

O-C diagrams are indeed an important tool for the determination of observed evolutionary changes in W UMa overcontact binary systems. The collection (data mining) of these diagrams, utilizing the internet, is a very easy way to access this information.

9. Acknowledgements

I would like to thank the late Lee Snyder, one of my mentors in eclipsing binary photometry, for conversations we had many years ago about W UMa’s evolution and several talks and discussions at past Society of Astronomical Sciences Symposims on data mining which ultimately inspired this paper.

10. References

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12 W Ursae Majoris Variable; http://www.natureillustratedmagazine.com/q/W_Urse_Majoris_variable


O-C Diagrams for W Ursae Majoris Systems Type W

- W UMa
- AB And
- GZ And
- SS Ari
- 44 Boo
- AC Boo
O-C Diagrams for W Ursae Majoris Systems Type A

AH Vir

OO Aql

V417 Aql

S Ant

V535 Ara
Photometry on MOTESS-GNAT Variable Star Candidates

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Abstract
This paper provides results from follow up photometry on MOTESS GNAT 1 variable star candidates at SSC Observatories. Targets were selected that had short periods more likely to be subject to aliasing, given the one synodic day cadence of the MOTESS asteroid search observations from which the catalog was abstracted. Results for 14 stars are presented with several partial models of eclipsing binary systems detected. Aliasing problems with under sampled observations and the opportunities for follow up with smaller telescopes are discussed as they pertain to MOTESS and the pending results from new survey telescopes that are coming on line.

1. Introduction

In November of 2015, SSC Observatories’ automated observatory at the Center For Solar Systems Studies became fully operational, Stephens (2015). The observatory contains two telescopes optimized for photometric studies. As one of its initial programs at its CS3-7 facility, SSCO started a program to photometrically measure, verify and categorize selected stars from the MOTESS-GNAT 1 Variable Star Catalog, Kraus et al. (2007).

MOTESS is the Moving Object Transit Event Survey System, Tucker (2007). It is a drift scan system of three telescopes that operated from 2001 through 2007 scanning a band of sky 48.3' high at a Declination of +3.0 degrees. The initial program focused on identifying moving objects to discover minor planets. Subsequently, the data from the survey was reanalyzed looking for variable stars. The resulting MOTESS-GNAT 1 catalog contains 26,041 candidate variable stars.

2. Target Selection

SSCO’s program has initially focused on under sampled candidate stars from the catalog. The cadence of MOTESS is one synodic day. Nyquist’s sampling theory states that, for sinusoidal varying sources, the minimum sampling rate required to determine the period of the source is two times the signal frequency. For non-sinusoidal sources, the waveform can be considered as a sum of the fundamental frequency plus the weight sum of its phase shifted harmonic overtones. For complex waveforms, the sampling rate must be double the frequency of the highest harmonic overtone required to reconstruct the input signal.

This implies that no period less than two days is likely to be correct in the survey. Furthermore, impulse like transients or saw tooth like pulsating variables are likely to be incorrect even with ascribed periods of 16 or more days.

Initial target stars were selected to satisfy the requirement that they were visible at air mass of two or less from horizon to horizon for at least a week. Additionally, they must have MOTESS-GNAT ascribed periods of less than 90 hours. Finally, their magnitude must be less than about 16.5 so that they have good signal to noise ratios through standard V and R filters on the CS3-7 0.25m telescope. Fields that contained multiple candidates satisfying these criteria were given preferential treatment during the selection process.

Table 1 shows the list of MOTESS-GNAT variable stars candidates observed between November 2015 and April 2016.

Table 1: Observed MOTESS-GNAT Stars

<table>
<thead>
<tr>
<th></th>
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3. Observation and Reduction

Observations were performed autonomously using the CS3-7 0.25m telescope and the LCRO 0.30m telescope. Depending on the target and Moon brightness, exposure times ranged from 120 seconds to 360 seconds. Exposures were alternated between V and R band photometric filters. At the end of each session, the results files are automatically uploaded to a database server at SSCO’s office for reduction and analysis.

The observational data is calibrated and, if necessary, rotated by a PYRAF script. Photometric measurements are performed using MPO Canopus software. Once, several nights of data are accumulated, period analysis and plotting is performed using Canopus.

4. Results

Table 2: Summary of Results

<table>
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<tr>
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Figure 1 clearly shows MG1-42996 to be a detached Algol-type detached eclipsing binary system. The system has been modeled, but without radial velocity data and full coverage in R band photometry, the derived model is not compelling.

Figure 2 shows that MG1-83320 has no significant variability during the period indicated in the MG1 catalog. It is likely that this star is near the saturation level of the survey and that night to night variations in seeing and opacity modulated some pixels in and out of saturation, leading to a false positive in the MOTES catalog.

The fact that the star was among the brightest in the catalog and that the catalog flagged the star with a 1.00 probability of being a false detection indicates that similar targets are best ignored.

Figure 3 shows MG1-83323, another star near the survey’s saturation limit. Due to the 16 hour catalog period, without coverage from different longitudes, it was difficult to get period coverage for this star. Its 16 hour catalog period and the level of variation is far below that set in the catalog. Its being an eclipsing system is still possible. It will require
full period of coverage to fully retire it from the candidate list.

Figure 3: MG1-83323 V Band

Figure 4: MG1-84399 V Band

Figure 4 shows MG1-84399 data folded on its published catalog period. The graph indicates that this candidate is likely a false detection. Our V-R color index is approximately 0.4 making it an F4 star.

Figure 5: MG1-84683 V and R Bands

Figure 5 shows MG1-84683 to be an over contact eclipsing binary. Good full period coverage offered the chance to model the system. Figure 6 shows the resulting model created with Nightfall, Wichmann (2011).

Figure 6: WD Model of MG1-84683

The system mass ratio is near 1.0 with both Roche lobes filled to a factor of 1.30. The primary and secondary exhibits a temperature of 5889K. The system, as seen from Earth, is inclined at 62.68 degrees. The addition of radial velocity data would allow the determination of the component masses, the absolute luminosity of the system and its distance.
Figure 7: MG1-214281 R and V Band Data

Figure 7 shows MG1-214281 to be another over contact eclipsing binary system. A quick fit of the system yields a mass ratio of 2.3348, an inclination angle of 68.57 degrees, Roche lobe fill factors for both primary and secondary of 1.3, and effective temperatures of the primary and secondary of 5433K and 5754K respectively. Figure 8 shows the system as modeled.

Figure 8: Model of MG1-214281

MG1-215228 has a noisier data set than most others due to being a bit under exposed. Figure 9 shows the R and V band scatter is higher than the pervious examples. The best fit to the data is a detached eclipsing binary system with a slightly elliptical orbit. It implies that the fill of the Roche lobes will oscillate throughout the orbital period. The system appears to have a mass ratio of 0.85, an inclination of 56.07 degrees, primary and secondary fill factors of 0.989 and 0.998, an eccentricity of 0.052 with periastron at 185 degrees and temperatures of 4468K and 4579K. A model of the system appears in Figure 10.
Full coverage of MG1-215228 was not obtained. With partial phase coverage there are at least two different interpretations that can be placed on the data without any good argument for one over the other. In short we need “More Data”!

Figure 11 shows the current data plotted against a period of 23.1 hours. This period favors a moderate amplitude pulsating star. Unfortunately, with a period close to 24 hours, without extra longitude coverage, it will take nearly a month of telescope time at the next opposition to resolve the question.

Figure 11: MG1-472598 Assuming 23.1 Hour Period

Figure 12 shows the same data assuming a 45.003 hour period. In this case it appears the system may be an eclipsing binary.

Figure 12: MG1-472598 Assuming 45.003 Hour Period

Figure 13 shows MG1-472914 to be near contact binary system with a period of 6.82 hours, a mass ratio of 0.6148 and fill factors of 9.89 and 9.98. Given the precision of the modeling and data, this may be a contact system, or soon will be. For that reason, it may be a system meriting monitoring as it evolves.
Figure 13: MG1-472914 in V and R Bands

Figure 14 shows a model of the system.

Figure 15 shows MG1-473365 V Band data plotted against its catalog period. It shows significant variation, but below the published amplitude.

Figure 16 shows the same star plotted against a period of 11.886 hours. This pattern suggests it is possibly a pulsating variable.

Finally, Figure 17 plots the star against a 23.78 hour period, suggesting the possibility that it too is an eclipsing binary. With candidate periods of near 12 and 24 hours, it will take a coordinated campaign to settle the nature of this star.
Figure 18 shows MG1-584533 is a detached eclipsing binary star system with a period of 22.496 hours.

Figure 19 shows this system modeled with a mass ratio of 1.93, an inclination of 87.29 degrees with respect to Earth and fill factors of 0.500 and 0.516.
Figure 20 shows MG1-604023 to be a contact eclipsing binary star system with a period of 6.424 hours.

Figure 21 shows the system modeled with a mass ratio of 1.0032, an inclination of 71.30 degrees with respect to Earth and fill factors of 1.0 and 1.0.

Figure 22 shows MG1-670506 band data plotted against its catalog period. It shows small variations, but not suggestive of a clear period. Examination of the target star, shown in Figure 23, shows it is near a bright saturated star.

Figure 23 shows MG1-670506 and its bright companion.

It is likely that seeing variations, and air mass changes modulate the skirts of the bright companion’s PSF, contaminating measurements made using aperture photometry of this target. When time permits, the target will be re-analyzed with point spread fitting to see if variability persists with this form of analysis.

Figure 24 shows MG1-670509 to be a detached eclipsing binary system with a period of 21.586 hours.

Figure 25 shows the system modeled with a mass ratio of 5.076, an inclination of 80.61 degrees with respect to Earth and fill factors of 0.83497 and 0.2876.
5. Conclusions

Of the 14 stars examined, the majority of them are correctly recognized by the survey as variable stars. The selection of stars with badly under sampled periods created a strong selection bias toward eclipsing binary systems, and a weaker selection bias towards low amplitude pulsating stars. Although no pulsating stars are conclusively detected, a couple of possible candidates are presented for follow up observation.

In the next year, the target selection will be expanded to candidates with longer catalog periods. Stars with catalog periods in the range of hours to 20 days will be examined. With the wider range of catalog periods it is expected that more pulsating and other variable forms will be discovered.

The MOTESS-GNAT catalog provides an excellent opportunity to discover and characterize many new variable stars using small telescopes. Working with MOTESS-GNAT also gives a taste of what to expect when new large survey telescopes start to come on-line. These instruments are going to cover the whole sky, not a degree of declination. By simple extrapolation, they are going to produce more than 3x10^7 candidate variables down to magnitude 18 and many times more as they go deeper.

Just like MOTESS-GNAT, the large surveys are going to have difficulty accurately characterizing under sampled targets. Current plans have survey telescopes with cadences that cover the sky once to several times a week.

Targets with periods out to a month will likely be aliased and incorrectly categorized. The volume of data from surveys requires that target identification will be automated. It is likely that target identification will improve with experience, thus initial target lists are likely to miss certain classes of objects and have more false detentions. The volume of survey image data, image storage requirements and bandwidth necessary to distribute and download images will probably make reanalysis of survey detentions difficult without follow on observation with dedicated instruments.

For those with the desire to follow up, targets will be plentiful and the most efficient tool will be your own telescope. It allows you to dwell with high cadences on targets of interest.

The coming torrent of candidate targets has begun with Panstarrs, Kepler and will increase with TESS, LSST and VLT Survey Telescope coming on line. They will create a wealth of catalogs and targets for small telescope science to explore.

6. References


New Observations of the Variable Star NGC 6779 V6

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Abstract
Existing and new observations of the variable NGC 6779 V6, are examined and analyzed. While a few studies of the longer period variable Star V6 in the globular cluster NGC 6779 occur in the literature, there is some disagreement as to whether this star can be classified as a regular period variable (e.g. RV Tauri), or whether it is an irregular variable. New BVRI observations of the variable V6 were made over a multi-year period and analyzed. These new observations, when combined with other published observational data, suggest that this the variable, should be classified as an RV Tauri type star

1. Introduction

NGC6779, or M56 is surprisingly poorly studied, probably because it is close to the Galactic plane (l = 62.66°, b = +8.34°), Rosino (1951), although the cluster is readily observable from northern latitudes for more than five months of the year, at relatively high elevation angles. The literature lists this object as a globular cluster of low metallicity, visual magnitude 8.4, and an integrated $B-V$ color index of 0.86. It has an angular diameter of approximately 8.8 arc minutes and is at a distance of 32,900 light-years. NGC 6779 has an average magnitude of 15.31 for the 25 brightest stars, an overall spectral type of F5 (Wehlau and Hogg 1985). NGC 6779 is shown in Figure 1.

Figure 1: NCG 6779 from 2016

Very early work by Helen Davis (1917) and by Harlow Shapley (1920) indicated only two confirmed variables in the cluster but later observations (Sawyer 1939; Rosino 1944; Wehlau and Hogg 1985) show that at least eleven variables are in or near the cluster, with an additional star (V2) whose variability remains in question (Russeva 1999). These studies were primarily short-term photographic examinations of the cluster and the results are somewhat puzzling, in that typically there should be many more variables in a globular cluster such as this. Table 1 shows an abbreviated table of the variables listed by Wehlau and Hogg (1985) and Clement, et al, 2001.

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</table>

Table 1: NGC 6779 variables from Wehlau & Hogg (1985)

The variable M56 V6 is about 0.6 arcmin north of the center of M56, which a bright star but with close companions, as would be expected in a cluster. Samus, et al, (2009) lists coordinates of the variable as RA: 19 16 35.78, Dec: +30 11 38.8, J2000.

Wehlau and Hogg (1985), initially classified the variable as a probable RV Tauri type, with a period of 90 days. Ivanov, Russeva, et al, 2000 did a photographic study on the cluster, and also commented on the variable as being an RV Tauri.
type. Pietrukowicz et al., (2008), studied the variables for M56 in 2002-2004, and reported V6 as a RV Tauri star specifically due to the characteristic appearance of its light curve. Russell (1997) produced a spectrographic analysis of RV Tauri stars in globular clusters and indicated that M56 V6 is “almost certainly of the RV Tauri class,” but also said it was “the most puzzling star of the sample studied.” Just a year later, Zsoldos (1998) wrote a paper concerning RV Tauri stars in globular clusters and suggested there were no such stars, in NGC 6779, due to irregularities in the power spectrum from period analysis.

Based on this ambiguity in the published record on M56_V6, and varying criteria for inclusion in the RV Tauri class, an observing program was started in 2012 to observe this variable using modern CCD photometric equipment over a long period of time. New CCD observations of NGC 6779 would also provide other researchers with the imagery to find other variables in this cluster.

2. RV Tauri Stars

RV Tauri stars are known to be giant and supergiant stars, having masses close to that the Sun, spectral types at maximum ranging from early F to late G or early K, and most have low metal abundances. RV Tauri stars occupy the upper right end of the Instability Strip on the HR Diagram. They are known to be strong infrared sources, and some observations have shown that RV Tauri stars have substantial amounts of dust surrounding them (Wallerstein, 2002).

RV Tauri stars in globular clusters are quite rare, with only six known examples: M2, M5, M10, M2, M56, and Omega Centauri, Zoldos, (1998).

There are a number of published papers on RV Tauri stars and their respective behavior. RV Tauri stars have two classifications: one photometric, and one spectroscopic.

In general, the visual light curves of RV Tauri stars have alternating deep and shallow minima, and have periods between 30 and 150 days (Sterken and Jaschek 1996). Pollard et al. (1996) list a number of photometric characteristics for determining a star’s inclusion in the RV Tauri class of variable stars. While Pollard et al. list a total of nine characteristics, some are only applicable to their comparison of a group of RV Tauri stars. Hence, the applicable photometric characteristics to a single star, and this paper, can be summarized as follows:

a) There are alternating deep and shallow minima in the light and color curves;
b) The secondary minima depths are more variable than primary minima depths;
c) A mean phase lag exists between the color index and light curves;
d) During extremely deep pulsations, the photometric colors get very blue;
e) The shorter period RVa subclass exhibits a constant mean magnitude.

From their spectra, RV Tauri stars are classified as A, B, or C depending upon both their temperature, and the presence and strength of specific molecular lines in their spectra. Specifically, Russell, (1997) indicated that this spectrographic criteria should include metal to iron (e.g. Ti/Fe or Cr/Fe) abundances.

3. Observations

3.1 Instrumentation

Two telescopes were used to obtain data on this variable star. The author utilized a Meade LX200 10-inch (0.25m) telescope, with a Starlight Xpress MX 716 CCD camera, which has a 550 × 720 pixel array. One-minute exposures were taken at an f-ratio of 6.3 (for an effective field of view of approximately 11 × 14 arc minutes) with standard BVRI photometric filters.

The majority of images were taken with the AAVSOnet 20 inch f/4 folded reflector in Sonita, Arizona. This scope is equipped with a SBIG STL 6303 camera, with Johnson-Cousins B, V, Ic and Sloan r’ filters. Images were exposed for three minutes. The effective field of view is 47 x 31 arc minutes.

For the images taken with the 0.25m telescope, the images were processed by standard techniques with Maxim DL. For the 0.5 telescope, the images were processed by the AAVSO processing pipeline. All images were then uploaded to the AAVSO Vphot website for photometry processing.

3.2 Data Collection

M56 V6, was observed from March 2012 (HJD 2456010.91228) through March 2016 (HJD 2457433.03205). The photometric sequence was developed from an AAVSO finder chart with the variable the photometric sequence shown in Figure 2. The average photometric error for these observations was 0.05 magnitude in the V band. The average air mass for this series of observations was 1.39 or with an average zenith angle of approximately 44 degrees.
3.3 Light Curves and Color Indices

BVRI light curves from 2012 are shown in Figures 3 through 6. The overall variation for the B band is approximately 1.2 B magnitude, and 0.8 for the other light curves. Figure 7 shows a combined light curve.

While the minima are fairly well aligned among the V and R band color curves, the minima occur at different times for the B and I bands. The B band maxima being slightly ahead of the V or R band minima, and the I-band minima being slightly behind.

The Color Indexes from the 2012 observations are shown in Figure 8. Examination of color indices are rather flat, less than 0.3 magnitude, with only the B-V color indexes showing a variation of approximately 0.5 magnitude.
3.4 Observed Minima

The longer scale variation of the M56 V6 variable is seen in Figure 9, which is a plot of the V band magnitude variation from the time period from 2012 through 2016. Note that this figure uses a modified Julian Date (MJD) of the actual JD minus 2455000.

Examination of the light curves in Figure 9 shows the maximum magnitude observed during this period was 12.22 V mag, with the minimum being 13.31 V mag, for a total spread of 1.08 V mag.

Figure also show some of RV Tauri-type alternating deep and shallow minima in their V band light curves, (Pollard, et al, 1996). For M56 V6, this effect is not as pronounced as in other RV Tauri variables, Pollard, et al. (1996). Comparing the minima at MJD 6081 (13.12 V mag), MJD 6211 (13.05 V mag), and MJD 6352 (13.09 V mag), or MJD 6441 (13.26 V mag), it is obvious that the variation in minima is less than 0.2 V mag.

4. Analysis

4.1 Period Determination

Because of the alternating shallow and deep minima of RV Tauri stars, the usual method of determining the period from adjacent minima is not used. The established convention for these types stars places the primary or deepest minima at phase 0.0, and the secondary, or shallow minimum at phase 0.5. Additionally, the primary or fundamental period is defined to be the time between the deepest minima. The period between adjacent minima is then designated as the half-period, (Fokin, 1994). Complicating the period determination problem is the issue that it is typical for RV Tauri stars to generally show random fluctuations in period from one cycle to the next, (Percy, et al, 1997).

Using the AAVSO VSTAR application for period analysis initially yields a period of 44.37 +/- 0.02 days, and is seen in Figure 10. Using the preceding description of period determination for RV Tauri stars, this half-period is doubled to yield a period for the V6 variable as 88.74 +/- 0.02 days. This compares favorably with Pietrukowicz, et al (2008) of 89.7 days, with some allowances for random fluctuations in period.

Further examination of this period graph also shows the variations in the depth of the minima over multiple cycles in the width of the period plot.

5. RV Tauri Evaluation

Using the photometric criteria from Section 3 on the data for M56_V6, we find:

a. Alternating deeper and shallow minima in the light curves, although the variation is not very pronounced. Similarly, the color indices show even a more shallow variation in the depth of the minima
b. Due to the limited variation in deep and shallow minima and some sparseness in the data, it is not possible to determine if the secondary minima depths are more variable than primary minima depths
c. A small mean phase lag exists between the color index and light curves. Close examination of the color indices and the light curves shows on
average the difference is color index minima and the light curve is five days or 0.06 phase difference.

d. During a relatively deep pulsations near HJD at 2456078.85177, the photometric color index B-V was 1.36 which was the maximum index value during all of the 2012 observations.

e. As M56_V6 is part of the shorter period RVa subclass, it also exhibits a constant mean magnitude around its maxima and minima. Over the course of the 2012 to 2016 observing period the mean V magnitude was 12.70. The variation from the mean over this period from each observed cycle was 0.0004 V mag.

If we consider the available spectrographic analysis done on V6, Russell (1997) notes that spectra of this variable shows excellent agreement for Cr, Mn, and Ni abundances with other RV Tauri stars. However, four spectral lines of iron, and several other elements show abundances closer to spectral studies of red giant stars.

6. Conclusion

Given the RV Tauri photometric criteria outlined by Pollard, et al, 1996, M56_V6 generally matches the photometric criteria for an RV Tauri star, although weakly for some of the criteria. Likewise, the spectrographic evidence is good in some areas and not in others.

As RV Tauri stars are certainly variable stars in transition from one state to another, Russell (1997) may be correct in that “perhaps the most significant conclusion to be drawn from the present work, is that RV Tau variables are not a homogeneous class of stars at all. It seems more likely that they include stars from several different stages of evolution.”

In summary then, M56_V6 can be classified as a RV Tauri star, but it is still somewhat unlike other stars in that variable class.

7. Acknowledgements

The author wishes to thank the AAVSO, its former director, Arne Hendon, and the staff at the Sonita Research Observatory for their support and assistance in gathering the data for this paper.

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Mixed-model Regression for Variable-star Photometry

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Abstract
Mixed-model regression, a recent advance from social-science statistics, applies directly to reducing one night’s photometric raw data, especially for variable stars in fields with multiple comparison stars. One regression model per filter/passband yields any or all of: transform values, extinction values, nightly zero-points, rapid zero-point fluctuations (“cirrus effect”), ensemble comparisons, vignette and gradient removal arising from incomplete flat-correction, check-star and target-star magnitudes, and specific indications of unusually large catalog magnitude errors. When images from several different fields of view are included, the models improve without complicating the calculations. The mixed-model approach is generally robust to outliers and missing data points, and it directly yields 14 diagnostic plots, used to monitor data set quality and/or residual systematic errors—these diagnostic plots may in fact turn out to be the prime advantage of this approach. Also presented is initial work on a split-annulus approach to sky background estimation, intended to address the sensitivity of photometric observations to noise within the sky-background annulus.

1. Introduction

Photometric reduction of amateur astronomic images often proceeds image-by-image, comparing the unknown target’s instrumental magnitude to those of one or more comparison stars (“comp stars”) in the same field of view. Zero-point and extinction values need not be explicitly determined, especially if comp stars are close to the target star in color.

Other reduction approaches work across numerous images, even to “all-sky photometry” with its strong assumptions about the constancy of instrument and especially of sky conditions, at least during “photometric” nights. In both approaches, transforms and perhaps other general corrections are made to target magnitude estimates before reporting.

The mixed-modeling approach proposed here borrows from both comparative and all-sky photometric reduction. However, rather than computing target magnitudes from instrumental magnitudes in a linear, deterministic manner, the present approach first models all comp-star data, and then applies this model to target-star and check-star data to extract best estimates of target and check magnitudes.

While the proposed approach also makes assumptions about sky, instrument, and comp-star properties, these assumptions are very flexible, in fact can include any level of complexity warranted by the quantity, diversity, and quality of instrumental magnitude data taken. The resulting workflow, based on mixed-model regression as a model for one night’s photometric behavior, has advantages in unifying automated data reduction and especially in affording diagnostic plots valuable to user monitoring of optical and model performance.

2. A Mixed-Model Approach

Mixed-model regression (Gelman and Hill, 2007) is an extension of standard multivariate regression. Familiar multivariate regression has only an intercept and “fixed effect” terms, where each regression parameter is fixed across all input data points and is also linear in the model (even when data terms are non-linear). By contrast, mixed-model regression adds “random effect” regression terms, whose parameter values are fixed only within each user-defined category of input data points but very probably differ between categories. These additional random effects were adopted by social sciences and agricultural statisticians to isolate unpredictable “random” effects such as gender, agricultural test plot, or climate type. Such categories can confound a study’s model without being important to the study’s goals. That is, random effect terms were added to remove systematic errors caused by data points’ membership in different categories. The idea is similar to classical Analysis of Variance but is more general.

In earlier work, the author found (Dose, 2014) that image-to-image fluctuations in photometric zero-
point constituted exactly such a confounding systematic error, one that could be largely removed by adding an image ID categorical variable to the data set and a random effect term to an otherwise standard photometric model (as in Warner, 2006). When multiple stars were measured in each image, that earlier work demonstrated removal of about 60-80 % of each of two types of systematic errors: (1) per-image errors due for example to exposure time jitter passing cirrus clouds, or dew forming on the optics; and (2) per-catalog-star errors due to minor errors in catalog or sequence magnitudes. Each systematic error is addressed by adding one categorical variable value (image ID and catalog-star ID, respectively) to each data point and one random-effect term.

3. Implementing the Model

The mixed-model workflow presented here is being built to manage the author’s variable star data, for reporting to the American Association of Variable Star Observers (AAVSO). One night’s data set typically comprises 1000-3000 comp-star observations and delivers 50-200 reportable magnitudes on 10-70 targets, typically Mira variables. The workflow is now almost entirely automated except for visual image inspection and detection and removal of outlier data points.

The present approach does not use a formula to calculate target-star or check-star magnitudes. Rather, the comp star data are modeled, and then check-star and target-star magnitudes are back-calculated from the model.

The first step models instrumental magnitudes of all the night’s comp stars, yielding one model per night, per filter. The model is refined if needed, and the comp-star data set is curated to get a final model fit to all the night’s comp-star data, as in Section 3.3 below. Modeling and curation are blind to any check-star or target-star data. Once the diagnostic plots are examined and the comp-star input set and models are settled upon, the comp-star input data are discarded.

The models are then applied to all target stars and check stars, filter by filter, as described in Section 3.4. In short, a temporary, hypothetical magnitude is assumed for all target stars and check stars, and then the model is applied to predict hypothetical instrumental magnitudes, which are used in turn to extract the best estimate of the stars’ time-dependent magnitudes.

Workflow stages are described in the following subsections. All software to date was written by the author in the R statistical language (R Project); all source code to date is available online (Dose, 2016).

3.1 Field-of-view (FOV) files

The current workflow’s FOV files summarize all required properties of comp stars, target stars, and optionally check stars within a given image field of view. They are prepared in advance of any reportable photometry on those fields of view and are used twice per observing night: once before observing to generate automated observing plans, and once after observing to automate processing of the night’s images as described below.

A given FOV file contains the relevant AAVSO “sequence” of stars close to a given target star. While one could assemble sequences directly from tabular data, much better practice is first to take a deep (240-s or 480-s) image of the field of view in each photometric filter, build a sequence using VPhot (AAVSO) in the usual manner, then visually inspect the images, especially neighboring stars (which may appear in one filter but not in another). The curated VPhot sequence text is then safe to copy verbatim as a section of the new FOV file, to which are added: optimum image-center (RA, Dec), main-target ID and period of light fluctuation, AAVSO chart ID, locations (relative to given star) of any known interfering signals to be removed from sky-background annuli (“punches”, described in Section 6.1).

3.2 Preparing one night’s data

From each observing session, all image FITS files are subjected to consistency checks (e.g., name matching FITS OBJECT, proper plate solution, availability of a field-of-view file), and a backup copy is made of the uncalibrated images, used later to assess star saturation. Images are flat- and dark-calibrated in the usual manner and inspected visually.

A master data frame is assembled by an R script, using the FITS files and relevant FOV files. One subtlety is in choosing the archived, uncalibrated FITS files to automatically flag pixel saturation relevant to the associated calibrated image. The master data frame serves as the central data source for the rest of the workflow. Each row represents one star in one image, and each of 34 columns represents one data type, for example a unique serial number, star ID, applicable transform value, total ADUs, sky background and variance.

This master data frame, a native R data format, contains all the reduced data needed to build the mixed model, to extract and transform target magnitudes, and to automate writing of an extended-format AAVSO variable-star report. Once a night’s master data frame is assembled, its FITS files are no longer used.
3.3 Building the Model

The present photometric data-reduction approach is founded on a model built using all of one night’s comp-star instrumental magnitudes, one model per filter, one point per comp star. The simplest useful mixed-model formula (here, in the V filter) is:

\[ I_V = [M_V + T_{V,V-I} * C_{V,I}] + A + (1 | IID) \]

where:

- \( I_V \) = Instrumental magnitude (time corrected)
- \( M_V \) = Catalog magnitude in V for comp star
- \( T_{V,V-I} \) = Transform in V, relative to V-I, for optics
- \( C_{V,I} \) = Catalog color index V-I for comp star
- \( A \) = Airmass for image containing comp star
- \( IID \) = Image ID (categorical)

and where in the simplest case, all terms within brackets are known in advance: \( M_V \) and \( C_{V,I} \) originate from catalog values; transform values \( T_{V,V-I} \) are typically measured on a previous night but may be obtained simply by rendering \( T_{V,V-I} \) an adjustable parameter rather than a constant. Airmass \( A \) is extracted from each image’s FITS header, and the Image ID (expressed in mixed-model formulas as 1 | IID) is arbitrary so long as it is unique to each image; in practice FITS file names suffice as values. The regression intercepts, implicit but included, correspond to the (per-filter) zero-points of classical photometric methods. The intercept and IID terms are linearly dependent; this degeneracy is resolved by a convention that random-effect terms like IID are zero-centered.

The mixed-model regression and model predictions are implemented as calls from the author’s R scripts to the “lmer” and “predict” functions within the open-source “lme4” package (Bates et al, 2015).

All catalogued stars within the field of view, except target and check stars, may be used to enhance the data set on which the night’s model is built. Inclusion of a few standard star fields—especially at diverse airmasses—helps to stabilize the night’s models and improves detection of systematic errors and individual outlier points. AAVSO sequences vary greatly in comp star quality, and observing at least a few AAVSO sequences with very high-quality comp stars largely serves the same purposes as does observing of standard fields.

The minimal model formula may be readily extended to model and remove additional suspected systematic errors. For example, a “Vignette” data field computed as each star profile’s squared distance from CCD center can be used to remove any consistent quadratic vignetting not removed by flat-field calibration. Such a model formula (in V filter) would be:

\[ I_V = [M_V + T_{V,V-I} * C_{V,I}] + A + V + (1 | IID) \]

where:

- \( V \) = Vignette term, in pixels\(^2\) from CCD center

Terms can be added to isolate linear response gradients and other consistent data artifacts. The added terms can have very high statistical power, given the large number of degrees of freedom afforded by the many comp-star data points constraining each model.

Only the night’s master data frame and models are needed for rest of the workflow.

3.4 Using the Model to Extract Target Magnitudes

Extracting best magnitude estimates for target and check stars begins with predicting their instrumental magnitudes from their inherent magnitudes—not the customary reverse. That is, the workflow’s next stage begins with assuming fictitious star magnitudes, and then applying the model to these fictitious magnitudes, along with other known star data (e.g., image ID, airmass) and model-specific data (e.g., transform values, extinction coefficients). These pro forma magnitude predictions are wrong, of course, but the difference between the measured and the predicted instrumental magnitudes are related to the image’s zero-point. In fact, if zero is adopted for all fictitious star magnitudes, the arithmetic simplifies so that the difference between measured instrumental magnitude and the model-predicted instrumental magnitude is itself the best estimate of untransformed target-star and check-star true magnitude.

Next, color-correcting to a “transformed” magnitude estimate for each target star in a given filter requires two values: (1) a transform value, as \( T_{V,V-I} \) in the model formula above; and (2) the color index corresponding to the same transform value, as \( C_{V,I} \) in the model above. While color values may already be known for a minority of variable stars, for most variable stars the color index is a key unknown of interest (and may well be time-dependent).

To avoid assuming the very color values that one is trying to determine, the present approach requires that at least one image be taken every night for every target in every color-index filter. Instrumental magnitudes are linear in transform values, so the software solves simultaneous equations to extract transformed magnitudes in both filters and the derived color index values.
For variable stars—especially for Miras and many other Long Period Variables—these color corrections can be significant. However, modern optics often yield very small transform values, often less than 0.04 in absolute magnitude, so that in practice, approximate color index values are quite sufficient. To maximize time-series observation time in the main filter, the author’s automated telescope is programmed to make only occasional color observations, such as I-filter images during a V-filter time series. It has proven quite satisfactory to spline-interpolate color index values measured every 20-30 minutes to yield color index values over time series of several hours’ duration.

3.5 Curating and Reporting.

The workflow makes no distinction between check stars and target stars until final report formatting. That is, while check-star catalog magnitudes are necessarily stored in the master data frame to detect outliers during final report curation, check-star catalog magnitudes are never used in the model or in magnitude predictions.

Transformed magnitudes for one night’s check stars and target stars are the workflow’s computational end point. An attempt is made to detect and remove questionable results, typically by re-inspecting FITS images (e.g., for cosmic ray hits), and then the AAVSO report is generated automatically and uploaded.

4. Diagnostics

The single most important advantage of all-night modeling by mixed-model regression may turn out not to be in accuracy or in automation, but in the diagnostic plots that emerge naturally from the unified input data frame and from the mixed models.

An oft-expressed concern about automated data reduction is that observers may trust “black box” algorithms too much and “never look at the data.” The current workflow is designed specifically to avoid such hazards (1) by requiring that images by inspected individually before inclusion, and especially (2) by generating numerous and diverse diagnostic plots, which if well-chosen are more illuminating and exacting than visual inspection of diverse raw data in quantity. This project’s current R scripts generate 14 high-resolution diagnostic plots from a model and master data frame (1000-2000 data points) in about 2 seconds. This section gives several examples.

4.1 Finding Outliers in Raw Data

Sorted regression residuals (observed-model) for the comparison stars are plotted vs normal-distribution quantiles to yield a classical “Q-Q” (quantile-quantile) plot, which serves to expose both systematically non-normal residual distributions and comp-star outliers. This is the present approach’s principal means of responsibly removing comp-star outliers, that is, in removing outliers that are not simply large in residual but rather that definitely deviate from overall distribution of residuals. Figure 2 gives a Q-Q plot for a well-behaved data set.

![Figure 2. Classical “Q-Q” plot for all one night’s comparison stars observed in the V filter. Q-Q plots are used to verify approximately Gaussian dispersion and to detect outliers.](image)

4.2 Monitoring Imager and Model Performance

The present approach isolates image-to-image variation in its own random effect (IID term in the model formula of Section 3.3), giving a quick look at imaging stability throughout a nightly session. Each image has its own value of IID in units of magnitude, and plotting these values vs image or vs time yields a “cirrus plot”.

Figure 1. First light curves extracted via the proposed mixed-model regression workflow.

Figure 2. Classical “Q-Q” plot for all one night’s comparison stars observed in the V filter. Q-Q plots are used to verify approximately Gaussian dispersion and to detect outliers.
Figure 3. A “cirrus plot”: image-to-image variation (model IID term) in average magnitude of comparison stars, where by convention zero corresponds to the nightly photometric zero-point.

Figure 3 indicates mostly small image-to-image variations. The larger variations early in the session (fractional JD ~ 0.5) were in twilight, and late variations are suspected of arising from dew forming on the SCT front plate. Even these variations are probably too small to detract from photometric quality—the author has removed up to 300-500 millimagnitudes of such IID-term noise from dew or clouds without detectable influence on check-star magnitude estimates.

Figure 3 also demonstrates steady performance of the optical system over several hours. The cirrus plot is the current workflow’s best single expression of optical system performance throughout a nightly session.

One may also plot residuals vs any number of other variables to monitor model performance, and especially to ensure that the model’s terms suffice to account for instrumental magnitudes. For example, Figure 4 demonstrates that airmass is properly accounted for, and specifically that the data do not indicate need for quadratic-term or other airmass corrections. Residuals appear higher in the few points taken at higher airmasses, as expected, but no slope or curvature is evident.

Other potential systematic errors arising from incomplete modeling can be assessed by specific plots. For example the “vignette plot” of Figure 5 rules out significant vignetting remaining in this night’s images.

4.3 For Other Systematic Errors
Diagnostic plots of model residuals can expose unexpected systematic errors. For example, Figure 6 indicates that a minimal model lacking sky-background correction terms has underestimated faint stars’ magnitudes or has overestimated bright stars’ magnitudes, or both. Sky-brightness estimation and modeling are covered in Section 6, below.
5. Evaluation of the Mixed-Model Approach

Strengths of this mixed-model approach and current implementation include:

- Very flexible modeling, especially in the ease of adding terms to account for various systematic errors. For example, current models are constant in extinction (per filter) through a night, but it would be straightforward to add a spline or other smoothed model for hourly extinction changes; the very numerous degrees of freedom would likely support any additional parameters needed.

- Partial removal of a few common systematic errors.

- Inherent consistency of parameter values, as their values all arise from the same model regressions;

- Statistical robustness to strongly imbalanced data sets, specifically to widely differing numbers of comp stars in a night’s images;

- Scalability to large data sets with little increase in user effort;

- Automated diagnostic plots, which aid both in setting the minimal model suitable to a given night’s data and in responsibly removing outlier data points.

Obstacles requiring more work (now in progress) include:

- The requirement for extensive custom software. While mixed-model regression and predictions require few lines of local code, data-management and photometric calculations have now grown to thousands of local code lines—far beyond R’s best use as a rapid prototyping framework. The complexity will probably require migration to a more modular, object-oriented language such as C#, Java, or Python.

- Sensitivity to low-quality comp stars. A model’s quality turns out to correspond to the lowest-quality comp stars included, not to their average quality. That is, the notable strength gained by accumulating very many model degrees of freedom can be dissipated by just a few bad input points. Unfortunately, comp star magnitudes vary widely in reported quality. To address this concern, work is in progress along 4 lines: (1) a tiered approach, running several models per night each with a subset of comp stars with different quality tolerances, to exclude low-quality comp stars from all model runs in which they are not needed; (2) weighting models’ input data points by reported quality; (3) including a per-catalog-star random effect term to isolate the worst comp-star errors; and (4) in separate observations, to rapidly cross-observe standard-field and low-quality comp stars to improve the FOV files’ comp-star catalog magnitudes.

- The power of mixed-model regression to isolate systematic errors has surfaced a small but consistent source of error: in the estimation of sky background fluxes by popular algorithms. First work on improving this situation is given in the next section.

6. Improving Sky-background Estimation

Early diagnostic plots from this work strongly suggested that photometric accuracy of the faintest stars was limited not by statistical “shot” noise, but by bias in estimating background sky brightness at the target signals.

Terrestrial observations of a star signal are the sum of its own (extinction-modified) “true” signal and an overlying diffuse signal from atmospheric light scattering. This sky-background signal is usually measured from the aggregate signal of pixels in an annulus outside the star signal but centered on it. This per-pixel signal is then subtracted from the pixels representing the target star, and this sky-corrected star signal is used to compute the star’s instrumental magnitude in that image.

Any bias in sky-background estimations results in a systematic error in derived instrumental magnitudes, and the effect is strongest for the faintest stars. To date, the author has isolated three relevant and very common problems: (1) leakage of target star flux into the sky-background annulus, (2) presence of flux sources (neighboring stars) in the annulus, and (3) the exact means of averaging the annulus pixel analog to digital units (ADUs)—mean, median, trimmed mean, etc.

The first problem—leakage of star flux into the sky-background annulus—is serious in effect but probably requires only a larger annulus inner diameter or an image-by-image adjustment to compensate for variations in star profile diameter between images. It is not considered here.

The author’s early algorithmic measures to address the other two difficulties with sky-background estimation are presented here.

6.1 Explicitly removing interfering star profiles

To deal with stars falling in the sky-background annulus, one can start with a robust means of aggregating the annulus pixels’ ADUs to a best estimate. Clearly, mean ADU is a very poor choice. Median ADU is a better and more popular choice, as it tends to reject all but the worst interferences.
Outlier pixels due to interfering stars in the annulus are far from randomly placed: indeed, for neighboring stars, the outlier pixel locations are maximally correlated. But location-independent aggregations like the median fail to take advantage of this signal correlation.

Probably the most obvious remedy is to remove affected annulus pixels from consideration. Since each interfering star’s location is practically constant, we record it as a new entry in the relevant FOV file. In effect, pixels affected by each interfering star are “punched” out of the annulus, as in Figures 7 and 8.

![Figure 7. Image detail around target star (at center) and brighter interfering star at lower left.](image7.png)

![Figure 8. Sky-background annulus (gray, outer ring) with “punch” effected over position of interfering star to remove its effect on median ADU.](image8.png)

Finding and recording the locations of interfering stars does require a one-time effort. First, preliminary images are taken, preferably at long exposures for best star detection, and then each interfering star location is recorded as a “punch” record in that star’s field of view (FOV) file. The author finds that recording interfering stars’ locations in arcseconds north and east from the target star performs better than does recording separate RA and Declination: relative locations probably tend to cancel the effects of imperfect star profiles or plate solutions. Recorded interfering stars are then removed automatically from every annulus, for every future photometric image.

### 6.2 Split-annulus sky-background estimation

For modern CCDs, pure sky-background noise is expected to be diffuse and largely independent of pixel location, but many other noise sources (other than the neighboring star) are highly localized, including:

- Asteroids moving through the field of view;
- Previously unrecognized, unremoved stars;
- Incomplete removal of interfering stars by the punch algorithm; and
- Cosmic ray hits or other CCD noise specific to one image.

The next improvement in sky-background estimation is built on a median-of-median algorithm, an extremely robust aggregation method in which one (1) divides the data set into roughly equal subsets, (2)
takes the median of each subset, and (3) takes the median of all the subset medians.

The key to success is in step (1), making subsets, and specifically in partitioning the pixels into subsets most likely to isolate correlated outliers into the smallest number of subsets. Happily, a naïve and easily automated approach—splitting the annulus into roughly equal slices—turns out to be nearly optimal.

![Image of sky-background annulus split into 12 equal slices and then "punched" to remove interfering star (same image detail as Figures X and Y).](image)

Figure 9. Sky-background annulus split into 12 equal slices and then "punched" to remove interfering star (same image detail as Figures X and Y).

In practice, choosing a number of slices between the cube root and the square root of the number of data values (annulus pixels) works well. For typical annulus geometries, this becomes 8-20 slices. Whenever punches are also involved as in Figure 9 and 11, it is best to remove slices that have lost more than perhaps half their pixels, keeping only the remaining slices for the final median.

In Figure 10, the interfering star at right is removed by the punch algorithm, but of course the transient cosmic ray hit seen below the target star is not. Still, Figure 11 suggests that the cosmic ray effect is well isolated into one annulus slice, minimizing its effect.

![Image of split-annulus approach isolates and minimizes the cosmic-ray hit's effect.](image)

Figure 10. Image detail around target star (at center), brighter interfering star at lower right, and cosmic-ray hit below the target star.

Figure 11. The punch algorithm removes the known star’s effect on estimated sky background flux, and the split-annulus approach isolates and minimizes the cosmic-ray hit’s effect.

Work continues on the statistical behavior of split-annulus aggregations, especially on whether median-of-median generally outperforms rival algorithms like trimmed-mean-of-medians, etc.

6.3 Modeling any remaining sky-background bias

Recent work confirms that it is feasible to model and isolate a general overestimation or underestimation of sky background flux across all of one night’s images. Unfortunately, the author has
also found that carelessly adding an ad hoc sky-flux correction term to the model can introduce its own serious bias, which in fact can be worse than those the term was intended to remove. The author is working to define the details both of sky-background estimation and of stable model correction terms to isolate and remove any remaining bias.

7. Equipment

Image data were taken with a Celestron Edge HD C-14 and matching focal reducer, PreciseParts custom spacer to achieve exact design backfocus, FLI Atlas focuser, SBIG STXL-6303E CCD at -35C with guider and filter wheel housing Astrodon V, R, and I filters, and Paramount MX+ mount. Panel flat frames are taken automatically at the end of most nights; dark frames are taken each 1-2 months.

This automated system is operated remotely, driven by Astronomer’s Control Panel (ACP) 8.0 on a laptop sealed within a small temperature- and humidity-controlled chamber permanently housed in a roll-off roof observatory in rural Kansas.

8. Acknowledgements

The R Project and R Foundation are gratefully acknowledged for making available without charge the rapid-development, open-source R language and packages on which this work was built. RStudio is acknowledged for making available the RStudio integrated development environment without charge. The AAVSO is acknowledged for making available photometric reference data and numerous web software tools that supports this work.

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Predicting a Luminous Red Nova

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Abstract

Luminous Red Novae (LRN) are rare transient events believed to be caused by the merger of a main sequence contact binary. Since the discovery of the prototype, V838 Mon, only a handful of LRN events have been observed. Tylenda et al. (2011) analyzed the OGLE data preceding the 2008 Novae of V1309 Sco and found that it exhibited a similar light curve to that of a contact binary with one interesting exception, the orbital period of V1309 Sco showed exponential period change going to zero. Unfortunately the system was discovered to be a binary after the merger, preventing any targeted observations to narrow down how the system entered this unusual state. However the extreme period change observed in V1309 Sco gives us a signature to look for in other contact binaries, allowing the discovery of merger candidates for follow up.

We will present an analysis of light curves and spectra of KIC 9832227 (NSVS 5597755) that show it is a contact binary system with a negative period derivative that is becoming more extreme with time. These data span more than 15 years and are taken from the NSVS, ASAS, WASP, and Kepler surveys, with ongoing measurements from the Calvin College Observatory and the Apache Point Observatory. The ongoing period change observed in the system is consistent with the exponential model fit from V1309 Sco and the rate of period change has surpassed that of all other measured contact binaries with the exception of V1309 Sco. If the exponential period decay continues the system will likely merge between 2019 and 2022 resulting in a naked eye nova. If this event occurs, this star will present the unprecedented opportunity to study a LRN progenitor and to follow the evolution of the merger.

1. Introduction

Luminous Red Novae (LRN) or V838 Mon type eruptions have been an area of considerable interest since the eruption of V838 Mon in 2002 (Munari et al. 2002). Several additional outbursts were then added to the new nova type. These include galactic LRNs V4332 Sgr (Martini et al. 1999) and V1309 Sco (Mason et al. 2010) as well as several extragalactic examples.

The observed LRN have typically shown a rise in brightness of ~10 M, over the course of about one year followed by a decay over the next several years. In contrast to more conventional novae LRN spectra evolve to that of a K to early M type as the novae
progress and end with the spectrum of a late M type (Mason et al. 2010). In addition numerous “bumps” of increasing luminosity are often seen in the downward decay. LRN are believed to occur within our galaxy at a rate of 0.1 - 0.5 year$^{-1}$ (Kochanek et al. 2014).

Numerous models to explain these atypical novae have been proposed. Fortunately the unique case of V1309 Sco was able to eliminate all models save for the binary merger model presented in Soker & Tylenda 2003. In addition V1309 Sco provides a period signature that can be used to discover LRN progenitor candidates years before the stellar merger occurs.

We will show how properly applying this template to public survey data can reveal candidate progenitors years before they enter stellar merger. We will present the first of our merger candidates, KIC 9832227, and explore the system’s physical parameters. And we will show how to eliminate the known sources of false positives with follow up observations.

2. V1309 Sco

V1309 Sco was discovered as a nova in September of 2008 (Nakona 2008). Shortly after its discovery it was determined to be of the LRN class (Mason et al. 2010).

The system was located close to the Galactic center and was observed by the OGLE survey. More than 1300 measurements, in the $I$ Cousins band, were made of the system prior to the beginning of the nova in 2008 (Tylenda et al 2011). Tylenda was able to show that the system began as a contact binary (W Ursa Majoris class variable star) with an orbital period of $\sim$1.4 days (Tylenda et al 2011). Once the outburst had subsided the system no longer showed evidence of its binary nature, implying that a stellar merger had occurred.

Tylenda also discovered that the orbital period of V1309 Sco decreased by 1.2% with an increasing rate in the years before the merger. From this period change Tylenda derived the following equation:

$$ P = 1.4456 \exp \left( \frac{15.29}{t - t_0} \right) $$

where $P$ is the system’s period in days at the time of observation, $t$, and $t_0$ is the time of merger (Tylenda et al 2011). The other values are constants fit to the V1309 Sco case. This model is presented in figure 1.

![Figure 1: The orbital period of V1309 Sco with the best fit line as presented by Tylenda et al. 2011.](image)

Unfortunately the binary nature of V1309 Sco was discovered after the merger occurred. This prevented any targeted follow-up to determine the physical parameters of the system. Thus V1309 Sco was unable to explain why a contact binary star system, one of the most common types of variable stars, would suddenly merge. However the period signature could be potentially detected years before the merger occurs.

3. Detecting Period Change

Small scale period change has been observed in many contact binaries. Traditionally period change has been measured by fitting minima of the light curve with a curve to find the $t_0$. The times of minima and maxima can then be compared to the predicted times, as indicated by the ephemeris. This technique could be used to detect the merger signature; however it requires fairly complete lightcurves of contact systems spread across several years. In addition any change in the lightcurve, for example star spots, can swamp any evidence of period change.

Detecting merger candidates with traditional period analysis would require vast amounts of data. Tens of thousands of contact binaries would need to be monitored for years before there would be enough data to detect the merger signature. This also eliminates most full sky surveys as they tend to spread observations out with respect to time.

Rather than using minima fitting we use Fourier transform modeling of the entire lightcurve. Once the model has been created the timing of any part of the lightcurve can be measured. This has several advantages over conventional modeling. First this
method results in much better signal to noise ratio, allowing for earlier detection. Also it does not require complete maxima or minima, allowing existing surveys like NSVS, ASAS, OGLE, WASP etc. to be used. In addition, because the whole lightcurve is used, any change in the lightcurve has a much smaller effect.

Fourier timing measurements combined with existing publicly available surveys could allow for most of the known contact binaries to be searched for the period signature and then identify priority targets for further effort.

4. KIC 9832227

4.1 History

KIC 9832227 is a variable star in the constellation of Cygnus. It was first identified as a variable star by the Northern Sky Variability Survey (NSVS) in 1999. However even in 2013 the nature of the variability, specifically whether it was a pulsating variable or a contact binary, was uncertain (Kinemuchi 2013). The Authors first studied the system to determine if the system was a contact binary.

In the process of determining the type of variability it was discovered that the system’s current period was shorter than the reported period. The original survey data was then reanalyzed and the period change was confirmed. In the process several other data sets containing KIC 9832227 were found.

Once the data sets were standardized the period change was measured via the Fourier model. The prototype lightcurve was generated from the Kepler data and the standard period and $t_0$ was selected to fall in the middle of the available data. The resulting O-C diagram is shown in Figure 3. The measured period over this time is shown in Figure 4.

The O-C data through 2014 were fit with the equation derived from V1309 Sco (Molnar et al. 2015). Figure 3 shows the best fit exponential model. Data taken in 2015 (also plotted) followed the extrapolation of that fit with no change to the parameters. The fit indicates a time of merger between 2020 and 2022.

4.2 Triple System Model

The exponential period change signature could have also been explained by a triple-star system. If the binary was in an elliptical orbit around a third star the resulting change in distance could explain the resulting period change. Triple systems are fairly common. Observations show that more than 20% of binaries are in hierarchical triples (Rappaport et al. 2013).
Parameters of a triple system with the minimum third star mass needed to fit the KIC 9832227 period change are given in Table 2.

<table>
<thead>
<tr>
<th>Parameter</th>
<th>Value</th>
</tr>
</thead>
<tbody>
<tr>
<td>Mass Binary (M$_\odot$)</td>
<td>1.72</td>
</tr>
<tr>
<td>Mass Companion (M$_\odot$)</td>
<td>0.6</td>
</tr>
<tr>
<td>Orbital Period (Years)</td>
<td>22.5</td>
</tr>
<tr>
<td>Eccentricity</td>
<td>54</td>
</tr>
<tr>
<td>Temp. Binary (K)</td>
<td>5870</td>
</tr>
<tr>
<td>Temp. Third Star (K)</td>
<td>4000</td>
</tr>
</tbody>
</table>

Table 2: Possible triple model

This is a potentially common source of false positives. Obviously this model needs to be eliminated before a system can be considered to be a good merger candidate.

The three body model can be eliminated with several different approaches. First the system can be monitored until the period change reverses or the star merges. This guarantees that the period change was not caused by a three body system but also prevents the prediction of a LRN in advance. Also given the abundance of triple systems the technique could require the monitoring of a large number of systems.

Another approach is to monitor the period change until it reaches the point where a third star could not be concealed in the system’s color and temperature. As the rate of period change increases the required mass of the third star also increases. Eventually this means that the third star’s light and temperature contribution to the system can no longer be hidden. This requires careful analysis of the star’s light curve and colors. The more different in temperature the third body is the easier it is to separate from the primary binary.

By far the best approach is to spectroscopically analyze the system. This allows for the detection of any third star as it is stationary relative to the primary contact binary. The star will show up as a separate set of emission and absorption lines. Also the addition of a third body with a different temperature will distort the blackbody curve of the spectrum.

Spectroscopic data can also be used to more precisely model the contact binary’s physical parameters. In the case of KIC 9832227 spectra taken at Apache Point Observatory and the Wyoming Infrared Observatory in the summer of 2015 have effectively eliminated a third body. Figure 5 shows the broadening function of APO data relative to a 4000 K template done to search for the presence of a third body. Figure 6 shows a set of broadening functions of KIC 9832227 in the WIRO data relative to a 5750 K spectrum. It shows the signature of a binary, with two peaks shifting position with orbital phase.
4.3 The Physical Model of KIC 9832227

We have also generated a physical model of the contact binary (Table 3, Figure 7). The system appears to be a fairly typical contact binary. None of the system parameters indicate why the binary appears to be in the process of merging. Note that the errors given in Table 3 are statistical errors, not model errors. The errors are likely underestimated as there are several possible models that fit the data. Items without error are taken from standard tables. The models were generated with the spectral types and masses. They are generated in WDWint, a binary modeling program (Nelson, R. 2006).

```
Mass Primary (M☉)        1.4
Mass Secondary (M☉)      0.32
Mass Ratio               0.228(.002)
Eccentricity             0(0)
Temp. Primary (K)        5870(19)
Temp. Secondary (K)      5900(18)
Inclination              54.74(.08)
Omega Potentials (1&2)   2.247(.003)
Bolometric Albedo (1&2)  0.5
Gravity Darkening (1&2)  0.32
```

Table 3: The best fit model for KIC 9832227

5. Future Work and Conclusion

V1309 Sco and KIC 9832227 both show that it is possible to detect stellar merger period signatures before the merger occurs. Even if KIC 9832227 does
not merge it proves that the tools and techniques that we have developed will work to identify candidates. We need to systematically search through the available surveys to identify additional merger candidates. Then a systematic follow-up of possible targets can begin. This will require dense time domain data, making it the perfect project for small observatories.

Also KIC 9832227 needs to be monitored until the system merges or the period change reverses. This will allow for the evolution of the system to be studied and will hopefully lead to an understanding as to why binary star systems merge.

In addition to the orbital changes the system needs to be monitored for changes in the spectral type and physical parameters. If it continues to approach merger we would expect to see significant changes in the best fit model parameters.

6. Acknowledgments

Funding for this project was provided by the Michigan Space Grant Consortium through a NASA grant, the Integrated Science Research Institute and the Calvin College Physics and Astronomy department.

The Apache Point Observatory and WIRO for the spectral data.

Special thanks to Optec, Inc. for facilitating my presenting at this conference.

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Abstract
It has long been suspected that white dwarfs accrete asteroid debris as evidenced by heavy metals in many white dwarf spectra. WD1145 was initially detected in Kepler data as an exoplanet candidate with a repeating 1.3% dip over the course of the Jul-Sep 2014 observing season. Follow-up ground based observations were conducted with professional telescopes during March through May of 2015, and these showed that the Kepler dip must likely consist of deeper and shorter dips which come and go with an uncertain pattern. It was hypothesized that the observations were due to an asteroid in a 4.5 hour orbit. In anticipation of its return to nighttime visibility, major observatories scheduled time starting in 2016 Feb. A pro/am collaboration was formed in late 2015 for amateur observations prior to the 2016 Feb professional observations in order to determine an ephemeris for fade activity for the purpose of scheduling relatively short observations with professional telescopes. The amateur observations began in 2015 Nov, sooner than requested, and they showed that the fade activity level had exploded, becoming ~20 times the level measured by Kepler. As many as 13 different fades per 4.5-hour orbit were measured, and these varied in depth from night to night. The amateur project turned into a full assault on the star with as many as 4 amateur telescopes observing on the same night. Continuous monitoring mysteriously showed that the clouds drifted in phase with respect to the dominant period i.e., they have a shorter period than measured by Kepler; this would imply that the orbiting dust clouds were located inside the orbit of the parent planetesimal. The best model indicated that the parent planetesimal was releasing fragments from inside its Hill sphere at the L1 Lagrange point, causing them to fall into an inner orbit. New astrophysics was described for the first time when the team used the diameter of the planetesimal orbit, and the diameter of the drift fragment orbit, to calculate the diameter of the Hill sphere from which the mass of the still unseen planetesimal could be inferred. Additionally, retroactive plotting of the drifting fade events backward in time hinted at a convergence date sometime in 2015 Aug, suggesting that this is when fragments broke away from the planetesimal’s L1 end and began the dramatic rise in fade activity; this tentative scenario requires more observational confirmation. Since 2015 Dec the fade activity subsided to ~ 5 times the level Kepler observed, occurring just before the scheduled professional observations. These amateur observations remain the most comprehensive to date.
1. Introduction

Past spectroscopic observations of white dwarfs show heavy metals contaminating their surface (Gänsicke et al. 2006, Jura 2008). This led to the theory that these stars were accreting asteroids but no examples had been found with this process happening in real time. Kepler initially detected this star in 2014 as an exoplanet candidate, but multiple periods ruled out the possibility of a single exoplanet (Vanderburg et al. 2015). Spectroscopic observations revealed aluminum, calcium, magnesium, silicon, nickel and iron putting it squarely in the polluted category of white dwarfs (Vanderburg et al. 2015). These elements remain on the surface of the star for a short period, usually less than a million years indicating relatively recent activity.

Follow up ground-based observations in March through May 2015 quickly confirmed Kepler’s findings of a 4.5 hour major period. Some observations showed up to a 40% dip in the light curve over 5 minutes (Vanderburg et al. 2015). Additional spectroscopic observations revealed a dust disk with an infrared excess and circumstellar absorption lines with 300 km s⁻¹ widths (Xu et al. 2016). These observations peaked interest in the object as the season for the star ended.

2. Methods

Plans for observing this dwarf with several major observatories were in place before the next season, in 2016. Major campaigns were planned for March when the object would be an all-night target. One group of professional astronomers decided to request amateur observations that would begin just prior to the 2016 professional observations. The initial goal was to just get a feel for the activity level, and not necessarily to produce publishable data. The previous Vanderburg observations suggested up to 6 asteroids in orbit. The Hereford Arizona Observatory initial light curves showed a surprising 11 dips during an exploratory observation in November, well ahead of when amateur observations were needed; this indicated a large increase in the activity of the dips (Figure 1.). This dramatic enhancement of the dips prompted 3 more amateur observatories to join the team. On many nights, two or more telescopes were observing the target.

After several nights’ data were reduced, attempts were made to reconcile the 4.5 hour period with the Kepler (K2) data. A 4.5 hour period was present, but additional peaks muddied the periodogram. Additional confusion was added when the duration and depths of the dips changed on a nightly basis. The complex shapes of the dips, inferred to be produced by the overlap of multiple individual dips, as well as their irregular occurrence, made traditional lightcurve analysis difficult. Therefore, a simpler method for characterizing the dips was utilized, wherein a bar replaces the dip whose length represents the duration of the dip while its thickness represents the depth. Figure 2 shows an example of how this type of representation allows us to understand the temporal behavior of the dips.

3. Discussion

With this new plotting system in place, a final collection of 37 nights over two and a half months were stacked in a “waterfall plot” phase-folded on the 4.5004-hour period (Figure 3.). The final plot was able to track the evolution of 15 discrete transiting clouds. Previous hypotheses suggested that dust clouds in orbit around white dwarfs would sublimate away in days. The final data showed that in fact the dust clouds around WD 1145+17 lasted for weeks to months with occasional outbursts. Secondly, the plot revealed for the first time numerous objects orbiting with periods slightly shorter than that of the parent body, which was completely novel. The inner orbit of the fragments was 4.4928 hours, or a mean drift of 2.5 minutes per day with respect to the 4.5004-hour period.
Figure 3. Waterfall plot phased on the 4.5 hour period and shows evidence of dust clouds drifting in a smaller orbit. There is also evidence of cloud evolution over time.

Multiple attempts were made to model the system with little success. Finally, an unusual approach was taken wherein the “chunks” were released from an asteroid filling its Hill sphere, leaving from the first Lagrange point (L1). This model set the bodies in a mostly circular orbit just inside the progenitor and was the best match to the data. With this information on the orbital distance for the parent body, and that of the fragment body, it is possible to calculate the mass of the parent body at approximately 1/10 that of Ceres. It is believed that this is novel astrophysics where for the first time the mass of an unseen planetesimal has been calculated using bodies released from its L1 point.

4. Conclusions

Subsequent follow-up observations show that most of the 15 fragments tracked in this work have disappeared presumably to sublimation. This suggests that outbursts of activity are possible for this star with a cycle time on the order of months. The mass of the parent asteroid suggests this star could see outbursts for thousands of years making it an ideal candidate for non-professional follow up. Ideally, much could be learned from capturing an initial outburst of new activity. Extending the drift lines backwards in time may suggest a tantalizing possibility that they converge on an origin date.

Following an outburst through its complete evolution could confirm or refute a single origination point for an outburst and would be a major leap in our understanding of white dwarf asteroid models. This case perfectly illustrates the advantage of telescope time, over telescope size, the former being the most advantageous area for amateurs. The scientific skills of the professionals meshed with high level amateur observatories, brought this study to fruition and should be a fine exemplar for future collaborations. At the time of publication, this is still the best all around body of photometry on the source, probably even including K2.

5. Acknowledgements

The authors would like to thank Cheryl Healy and Beth Christie for their material support of this project.

6. References


Astronomical Instrumentation System Markup Language (AISML)

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Abstract
The Astronomical Instrumentation System Markup Language (AISML) is an Extensible Markup Language (XML) based file format for maintaining and exchanging information about astronomical instrumentation. The factors behind the need for an AISML are first discussed followed by the reasons why XML was chosen as the format. Next it's shown how XML also provides the framework for a more precise definition of an astronomical instrument and how these instruments can be combined to form an Astronomical Instrumentation System (AIS). AISML files for several instruments as well as one for a sample AIS are provided. The files demonstrate how AISML can be utilized for various tasks from web page generation and programming interface to instrument maintenance and quality management. The advantages of widespread adoption of AISML are discussed.

1. Introduction

Light measured from an astronomical source is acted upon by the several types of instrumentation used to acquire it. Data, or properties, about these instruments (e.g. focal ratios, encoder resolutions, filter band passes and sensor read noise) are necessary for a range of tasks including system documentation, image acquisition and analysis and operation and maintenance. While many astronomical images contain some of this data in their FITS header, it is both far from complete and too closely, well literally, coupled with a particular image. The complete range of instrumentation data are typically saved in several formats scattered across different locations, making it difficult to locate and access (Figure 1).

Occasionally data must be manually copied from one format to another causing unnecessary duplication and potentially creating a source of error. In many cases an individual or organization can accumulate a sizeable inventory of instrumentation compounding the above issues. Thus a need arises for a uniform, standardized method for storing the properties of these instruments to facilitate accessibility and simplify maintenance and sharing. To meet this need, the first step is determining the best document format to hold instrument data.

Figure 1 Information about astronomical instrumentation systems is scattered over several subjects and sources.

2. AISML

2.1 Selection of XML

The Astronomical Instrumentation System Markup Language (AISML) maps properties about an astronomical instrumentation system (AIS) into an Extensible Markup Language (XML) format. XML

1 It's useful if the reader has some basic familiarity with XML (which can be found from the usual sources on the internet).
was selected because, in short, it’s the worst format except for all the others. That XML is the worst derives in part from its advantage of simplicity. In essence, while a well-formed XML document is simple to create, making practical use of it typically isn’t. Making practical use of an XML document requires additional knowledge of APIs and other XML implementations that aren’t as simple to learn and use. For example, a stand-alone, well-formed XML document requires a corresponding style sheet transformation document to view it in a formatted fashion in a browser.

The advantages of using XML that makes it a better format than the others are that it’s:

- A free, open and mature standard supported by the W3C.
- Easily customizable and, well, extensible.
- A text based format that’s both human and machine readable.
- Wide cross platform support.
- Supported by several free applications and APIs.

These advantages (Figure 2) aide the goal of seeking a wide spread adoption of AISML that will promote the development of freely available applications that facilitate the information management of AIS’s across the wider astronomical community. A standardized AISML implementation for instrumentation becomes the basis for applications that can be shared by anyone creating their own AISML documents. For example, AISML users can use an off-the-shelf style sheet transformation to view their implementation as a formatted web page without bothering the create their own.

![Diagram of AISML workflow](image)

Figure 2. The Astronomical Instrumentation System Markup Language (AISML) is centered on XML. XML provides several resources that allow AISML files to be used for both configuration and documentation.

### 2.2 Design of AISML

XML supports a hierarchical approach of entities called ‘elements’ with the constraint that each well-formed XML file must contain a unique ‘root’ element which can become the parent to any number of child elements. A critical design decision for AISML then becomes at what level in the instrumentation hierarchy root elements should be defined. At too high a level the AISML file can become large and unwieldy with too many child elements making it difficult to maintain. Too low a level can lead to too many small files which present their own maintenance problem and can make instrumentation properties unnecessarily difficult to locate.

To accommodate these considerations, each AISML file’s root element will be defined as an instantiation of a unique ‘instrument’ in the AIS, or as an instantiation of an AIS containing AISML instruments. An initial set of six instruments will be proposed for the AISML standard. These instruments\(^2\) (Figure 3) are the SITE, MOUNT, OTA (optical tube assembly), FOCUSER, FILTER_WHEEL and CAMERA\(^3\). An important

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\(^2\) ‘Courier New’ font will be used to identify AISML instruments.

\(^3\) A Windows user might notice some similarity between AISML and the ASCOM instrument models.
A criterion for defining an instrument is interchangeability. That is, an instrument must interchangeably act either upon light emitted by an astronomical source or on another instrument. For example, a primary mirror wouldn’t qualify as an instrument because it’s not typically interchangeable with the other OTA light collecting components such as a secondary mirror. An OTA on the other hand can interchangeably rest on different types of MOUNTs.

A seventh ‘instrument’, an AIS, is defined as a root element (see Section 4) whose children are derived from the root elements of its constituent instruments.

The root elements instantiating each instrument contain little information in and of themselves. Most of the information about an instrument is contained in the child elements defined lower in the hierarchy. AISML establishes three general requirements for the child elements under each root. First, each child (property) of a given root (instrument) should have a tag (property name) that is common for the instrument type, yet contain a value that is unique for each root instance. For example, a FOCAL_LENGTH child element would be common across all OTA root elements and would typically contain a unique value (excluding, say, a manufacturer’s inventory of an identical instrument type). Second, child elements should not contain redundant information. An OTA with a FOCAL_LENGTH tag would include either FOCAL_RATIO or APERTURE_DIAMETER as an additional tag, but not both since the third value can be derived from the other two. This requirement minimizes the chance of conflicting information. The final requirement is to limit the number of required child elements to ensure wide applicability and facilitate implementation. As a starting point, AISML will base its required child elements on those values typically used for instrument control software configuration. In addition to software configuration, it should be clear these elements can be used for documentation as well (Figure 2). The examples below will show how the user can define additional child elements that can be used for a variety of purposes ranging from maintenance to manufacturing and quality control. If additional elements used for other purposes gain a certain level of popularity they can either be added to the standard or made into a commonly adapted extension, similar to the SBIG FITS extensions.

As noted above, the interchangeability requirement of AISML instruments allows them to easily be combined as siblings under an AIS parent. For example a basic AIS might consist of a SITE, MOUNT, OTA and CAMERA. An AIS can range from a DSLR on a tripod to the Hubble Space Telescope. In the first instance the SITE would contain information including geographic coordinates; the OTA would include information about the lens, the MOUNT would describe the tripod and finally, the CAMERA would include information about the DSLR. In the Hubble instance the SITE would include information about the satellite’s orbit. A CAMERA could, with some imagination, be defined as the eyepiece/observer combination that leads to a hand drawn image. An AIS associated with a ‘raw’ image from a sensor could provide information about the image otherwise available in a FITS header. Although combining similar AISML instruments in an inventory fashion is possible (e.g. a new root OTAS contains child elements of OTA), it should be noted that AISML is not intended as an instrument inventory system.

One main difference is that ASCOM instruments are defined by drivers that require both methods and properties. ASCOM thus merges AISML instruments without methods, such as the OTA with the MOUNT to create their ‘telescope’.

Software configuration is used in this context to identify data input to a software application that affects the application’s operation.
3. AISML Instruments

3.1 SITE

As an instrument, the function of the SITE is to position the MOUNT. The primary configuration property of the SITE thus becomes its location. An example of a basic AISML file for the SITE root element is given in Listing 1. The elements and attributes (in upper case in the listing) are mostly self-explanatory, except perhaps for the ID attribute which will be required for each AIMSL root element. This attribute is essentially required to uniquely identify each instance of an instrument.

[Refer to LISTING 1]

It’s clear that SITE contains unique information for each observing location. Observers with portable setups can create a SITE file for each observing site they visit. The SITE listing provides an example of the advantages of XML. It’s a human readable text file that’s easy to maintain and share. It’s a machine readable file that can be used by any application that can parse XML. If you double click the file in Listing 1 it should automatically open in a browser window.

Several commercial off the shelf (COTS) applications such as TheSkyX and MaxIm DL, and websites such as Heavens-Above require the information contained in SITE for their configuration. If these applications adopt a common format such as AISML then the user would be able to point each application to the same file, rather than manually input the information multiple times.

It wouldn’t be XML if it wasn’t extensible and obviously a SITE can contain a lot more information. Listing 2 includes several additional tags that would not be included as part of the initial AISML standard. These extensions have been added to include additional information useful for documentation purposes. While readable, the AISML SITE file is clearly not optimal for displaying the information.

[Refer to LISTING 2]

The second line in Listing 2 refers to another XML document (Listing 3) that uses XSLT to translate the AISML/XML file into a document suitable for browser rendering. What this means is that if you now ‘open’ the AISML document with a web browser the browser will use the XSLT document to render the AISML file as defined in the XSLT source (see Figures 4 and 5). The added extensions include additional information, not necessarily required for configuration, but useful for documentation purposes. Using XSLT not only renders the information contained in the file but also enables several additional features such as Google Maps to be brought into use.

[Refer to LISTING 3]
Figure 5. The enhanced AISML SITE instrument in Listing 2 rendered in a browser using the translation file in Listing 3.

Readers are encouraged to copy either of the AISML files (Listing 1 or 2) and edit it to their local specifications. They should then copy the XSL file (Listing 3), put it in the same directory and open their version of the AISML file in their browser. Ambitious readers can also edit both files to further customize their implementation. It follows that similar XSLT implementations can be realized for each of the other AISML root elements identified below.

3.2 OTA

As an instrument, the purpose of the OTA is to collect (typically bringing to focus) light emitted by an astronomical object. In general, an OTA as its name implies, consists of optical components mounted in a tube assembly. An example of an AISML file for the OTA root element is given in Listing 4. The listing includes both the AISML required elements and attributes as well as some user additions.

[Refer to LISTING 4]

Again, an initial goal of AISML is to keep the required set of elements under each root small enough to have the widest applicability while enabling extensions for various user requirements. The AISML required elements were derived from the MaxIm DL telescope input parameters. (The OBSCURATION tag is clever since it’s dimensionless and obviates a need for defining the secondary.) Note that the additional OPTIC elements contain children useful for maintenance (CLEAN_DATE and COAT_DATE). Another possibility is to include children with components of ISO 10110 which defines XML tags for the presentation of design and functional requirements for optical elements and systems in technical drawings used for manufacture and inspection.

Recall that Figure 2 shows how AISML files can be used both for documentation and configuration. While Listing 3 provides an example that uses AISML for documentation, Listing 5 provides a simple example of configuration, using JavaScript to calculate the focal ratio for an OTA element. As above, the reader is encouraged to copy these files locally, modify and run. Listing 6 accomplishes the same task in Python (v2.7). Note that both listings use the XML Document Object Model (DOM) API which is both cross-platform and language independent.

[Refer to LISTING 5]

[Refer to LISTING 6]

Thus far, two of the six AISML instrument root elements have been defined. In addition to listing the required child elements under each root, several extension child elements have also been demonstrated. The extension elements show how AISML files can be used for a variety of purposes. An XSLT example for the SITE root element was provided (Listing 3) to demonstrate how the information contained in AISML documents can be conveniently displayed on a web page. It follows that such a document, if shared, would obviate the need for others to duplicate the effort. For the OTA root element two programming examples, one in JavaScript and one in Python, were provided to demonstrate how conveniently AISML documents can be accessed for calculation and other configuration requirements. It should be clear at this point that information saved in an AISML standard format, leveraging the advantages of XML, can easily be accessed by several methods without having to make copies into other formats.

The remaining instruments typically include firmware and a computer interface. Thus many instrument control programs may obtain configuration information directly from the
instrument. The advantage of maintaining an AISML file in this instance is that it enables information to be obtained when the instrument is offline. While an AISML file storing configuration information enables the possibility of containing conflicting information with what’s in the firmware, it’s possible to avoid this by writing a utility that compares the firmware and AISML file values to ensure they match.

3.3 CAMERA

The purpose of the CAMERA is to measure the light emitted by an astronomical object. An example of an AISML file for the CAMERA root element is given in Listing 7. Required AISML elements initially are limited to SENSOR_TYPE, PIXEL_INFO, INTERFACE and at least one MODE, containing BIT_DEPTH, READ_NOISE and GAIN.

[Refer to LISTING 7]

3.4 MOUNT

The MOUNT is the instrument responsible for positioning the OTA (relative to its own position defined by the SITE). An example of an AISML file for the MOUNT root element is given in Listing 8. The only required AISML elements under MOUNT at this time are the TYPE and INTERFACE tags under CONTROLLER.

[Refer to LISTING 8]

3.5 FOCUSER

The FOCUSER is an instrument that acts on the light as it’s passed from the OTA to the CAMERA. An example of an AISML file for the FOCUSER is given in Listing 9. In this instance VENDOR and MODEL are optional.

[Refer to LISTING 9]

3.6 FILTER_WHEEL

The last instrument to be enumerated is the FILTER_WHEEL. The function of the FILTER_WHEEL is to position various filters in the path of light before it strikes the CAMERA sensor.

An example of an AISML file for the FILTER_WHEEL is given in Listing 10. Initially for AISML, the tags VENDOR and MODEL will be optional and only NAME will be required under each FILTER.

[Refer to LISTING 10]

4. Instrumentation Systems

4.1 Types of AIS’s

As stated earlier, instruments (or AISML root elements) are defined such that they can easily be combined to create higher level instrumentation systems. One of the original goals for AISML was to define a format to describe the properties of instruments used in creating an astronomical image, let’s call it an AISi. It’s clear that most of the instruments defined via AISML don’t necessarily have to be used for imaging. Other types of instrument systems, e.g. using a video camera (AISv), spectrometer (AISs) or photometer (AISp) can easily be defined by adding the appropriate instrument to the existing standard.

4.2 AIS as an Instrument

To accommodate AIS’s within the AISML framework the last instrument to be defined will be an AIS. To create an AIS:

- The AIS tag must be used as the root element.
- XInclude should be used to point to the appropriate AISML instrument files defining the particular AIS.
- Each included instrument be contained within the appropriate AIS_[instrument] tag. The reason for this is to enable multiple instruments of the same type in an AIS file.

An example of a very simple AISi is given in Listing 11. This AISi includes a CAMERA and an OTA such that the attributes and values from these elements can be used to calculate the pixel and sensor field of views as well as the theoretical resolution limits. Listing 12 augments the AISi in the previous listing with a guide scope setup, using the TYPE attribute to identify the purpose (“Primary” or “Guide”) for instruments of the same type.

[Refer to LISTING 11]

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5 A typical DSLR CAMERA will have a Bayer filter array glued on top of the sensor. In this instance a FILTER_WHEEL isn’t typically included as the SENSOR_TYPE tag in CAMERA describes the filter pattern.

6 Additional instruments aren’t limited to what replaces a CAMERA in the focal plane. They can also include off-axis guiders, field rotators, etc.
5. Conclusion

AISML has been introduced as an XML based format for storing data about astronomical instrumentation systems. Using a defined format to save information about several types of instruments improves data quality by facilitating access and maintenance and minimizing redundancy. A defined format also simplifies sharing data in the wider community who then provide a larger pool for developing support utilities, potentially creating a network effect. XML was selected as the data format due to its widespread adoption as a standard and the availability of several free support utilities.

AISML files are intended for both documentation and configuration. Documentation is AISML content that becomes formatted for human consumption either via screen or paper, and configuration is AISML content that is used for machine consumption, typically for calculation, or as some other entity relevant to the particular application.

AISML files for six basic astronomical instruments were provided. The number of required properties is limited in order to facilitate wider adoption. Several additional properties are included in the listings as examples of AISML’s extensibility allowing for user customization. Some of these properties can be added to the standard if there’s widespread adoption. Customization not only enables adding user specific properties to each instrument, but also enables adding additional instruments to the collection. AISML defines a seventh, ‘systems’ instrument which enables the combination of other AISML instruments into an integrated instrumentation system.

6. References

Harold, Elliotte, 1999, XML Bible (Foster City, CA: IDG Books Worldwide, Inc.).


Listing 1. AISML SITE.

```xml
<?xml version="1.0" encoding="utf-8"?>
<SITE ID="SITE001">
  <NAME>Tierra Del Sol Observatory</NAME>
  <LOCATION>
    <ALTITUDE UNITS="meters">1131</ALTITUDE>
    <LATITUDE>32.61383</LATITUDE>
    <LONGITUDE>-116.33202</LONGITUDE>
  </LOCATION>
</SITE>
```

Listing 2. AISML SITE with extensions.

```xml
<?xml version="1.0" encoding="utf-8"?>
<?xml-stylesheet type="text/xsl" href="Site.xsl"?>
<SITE ID="SITE001">
  <NAME>Tierra Del Sol Observatory</NAME>
  <LOCATION>
    <ALTITUDE UNITS="meters">1131</ALTITUDE>
    <LATITUDE>32.61383</LATITUDE>
    <LONGITUDE>-116.33202</LONGITUDE>
  </LOCATION>
  <WEBSITE>http://tierra-astro.org/tdsobservatory.html</WEBSITE>
  <MPC_CODE /> <!-- Minor Planet Center Observatory Code -->
  <CLIMATE>
    <CLITYPE>
      <CLINAME>Weather Station</CLINAME>
    </CLITYPE>
    <CLITYPE>
      <CLINAME>Clear Sky Chart</CLINAME>
      <CLILINK>http://www.cleardarksky.com/c/TierradelSolCAkey.html</CLILINK>
    </CLITYPE>
  </CLIMATE>
</SITE>
```
Listing 3. XLST file used to render AISML SITE files in HTML.

```xml
<?xml version="1.0" encoding="UTF-8"?>
<xsl:stylesheet version="1.0" xmlns:xsl="http://www.w3.org/1999/XSL/Transform">
  <xsl:template match="SITE">
    <html>
      <body>
        <h3><xsl:value-of select="NAME" /></h3>
        <xsl:variable name="latitude">
          <xsl:value-of select="LOCATION/LATITUDE" />
        </xsl:variable>
        <xsl:variable name="longitude">
          <xsl:value-of select="LOCATION/LONGITUDE" />
        </xsl:variable>
        <table>
          <tr>
            <td><b>Latitude</b></td>
            <td><b>Longitude</b></td>
            <td><b>Altitude</b></td>
          </tr>
          <tr>
            <td><xsl:copy-of select="$latitude" /></td>
            <td><xsl:copy-of select="$longitude" /></td>
            <td><xsl:value-of select="LOCATION/ALTITUDE" /></td>
          </tr>
        </table>
        <p><a href="http://maps.google.com/maps?z=12&t=k&q=loc:{$latitude}+{$longitude}">Google Satellite</a></p>
        <xsl:choose>
          <xsl:when test="string-length(ROOF_TYPE) != 0">
            <p><b>Rooftype: </b><xsl:value-of select="ROOF_TYPE" /></p>
          </xsl:when>
          <xsl:otherwise />
        </xsl:choose>
        <xsl:choose>
          <xsl:when test="string-length(MPC_CODE) != 0">
            <p><b>MPC Code: </b><xsl:value-of select="MPC_CODE" /></p>
          </xsl:when>
          <xsl:otherwise />
        </xsl:choose>
        <xsl:choose>
          <xsl:when test="string-length(WEBSITE) != 0">
            <xsl:variable name="website">
              <xsl:value-of select="WEBSITE" />
            </xsl:variable>
          </xsl:when>
          <xsl:otherwise />
        </xsl:choose>
      </body>
    </html>
  </xsl:template>
</xsl:stylesheet>
```
Listing 4. AISML OTA including extensions.

<?xml version="1.0" encoding="utf-8"?>
<OTA ID="TDSOTA001">
<TYPE>Ritchey-Chretien</TYPE>
<FOCAL_LENGTH UNITS="mm">4762</FOCAL_LENGTH>
<APERTURE UNITS="mm">597</APERTURE>
<OBSCURATION>50</OBSCURATION> <!-- percent of APERTURE -->
<!-- Below are user additions -->
<OPTIC TYPE="Mirror" POSITION="Primary">
<COATING MAIN="Al" OVER="SiO2" />
<CLEAN_DATE>2014-11-21</CLEAN_DATE> <!-- YYYY-MM-DD -->
<COAT_DATE>0000-00-00</COAT_DATE>
</OPTIC>
</OTA>
Listing 5. Javascript file used to read AISML OTA parameters.

```html
<html lang="en">
<head>
<title>Listing 4</title>
</head>
<body>
<script language="JavaScript">
// Create a connection to the file.
var Connect = new XMLHttpRequest();
// Define which file to open and
// send the request.
Connect.open("GET", "o_TDS24.xml", false);
Connect.setRequestHeader("Content-Type", "text/xml");
Connect.send(null);
// Place the response in an XML document.
var TheDocument = Connect.responseXML;
// Place the root node in an element.
var Telescope = TheDocument.childNodes[0];
// Read the child element values
var FocalLength = Telescope.getElementsByTagName("FOCAL_LENGTH");
var fl = parseFloat(FocalLength[0].textContent.toString());
var Aperture = Telescope.getElementsByTagName("APERTURE");
var ap = parseFloat(Aperture[0].textContent.toString());
// Calculate
var fr = fl/ap;
// Write the results to the page.
document.write("Focal length: " + fl + ", APerture: " + ap + ", Focal Ratio: " + fr.toFixed(2) + ")
</script>
</body>
</html>
```

Listing 6. Python file used to read AISML OTA parameters.

```python
from __future__ import print_function
import xml.dom.minidom

def Listings():
    DOMTree = xml.dom.minidom.parse("t_TDS24.xml")
    Telescope = DOMTree.documentElement
    Focal_Length = Telescope.getElementsByTagName("FOCAL_LENGTH")
    f = float(Focal_Length[0].childNodes[0].data)
    Aperture = Telescope.getElementsByTagName("APERTURE")
    ap = float(Aperture.childNodes[0].data)
    fr = f/ap
    print("Focal length: ", f)
    print("Aperture: ", ap)
    print("Focal Ratio: ", "{:2f}".format(fr))
if __name__ == "__main__":
```

Listing 7. AISML CAMERA with extensions
<?xml version="1.0" encoding="utf-8"?>
<CAMERA ID="MNZC001">
  <VENDOR>SBIG</VENDOR> <!-- bought out by Cyanogen -->
  <MODEL>ST-402ME</MODEL>
  <SERIAL>N/A</SERIAL>
  <COOLING FROM_AMBIENT="true" UNITS="C">-30</COOLING>
  <CHIP>
    <SENSOR_TYPE IS_CCD="true">0</SENSOR_TYPE> <!-- ASCOM SensorType value -->
    <VENDOR>KODAK</VENDOR>
    <MODEL>KAF-0402ME</MODEL>
  </CHIP>
  <INTERFACE ASCOM="true">USB 2.0</INTERFACE>
  <PIXEL_INFO>
    <NX>765</NX>
    <NY>510</NY>
    <SX UNITS="microns">9</SX>
    <SY UNITS="microns">9</SY>
  </PIXEL_INFO>
  <MODE> <!-- A MODE tag is required. Use an attribute for multiple modes -->
    <BIT_DEPTH>16</BIT_DEPTH>
    <LINEAR_FULL_WELL UNITS="e-">100,000</LINEAR_FULL_WELL>
    <READ_NOISE>
      <!-- Use XSD YYYY-MM-DD, zero date is spec sheet -->
      <DATE>0000-00-00</DATE>
      <!-- Multiply by gain for read noise in e- -->
      <VALUE UNITS="ADU">9.2</VALUE>
    </READ_NOISE>
    <GAIN>
      <DATE>0000-00-00</DATE>
      <VALUE UNITS="e-/ADU">1.5</VALUE>
    </GAIN>
  </MODE>
</CAMERA>

<?xml version="1.0" encoding="utf-8"?>
<MOUNT ID="TDSM001">
  <TYPE>German Equitorial</TYPE>
  <!-- Below are user additions -->
  <VENDOR />
  <MODEL />
  <PIER>Reinforced concrete</PIER>
  <CONTROLLER>
    <VENDOR>Sidereal Technology</VENDOR>
    <MODEL>Brushless Servo Controller</MODEL>
    <INTERFACE ASCOM="true">USB 2.0</INTERFACE>
    <AUTOGUIDER_PORT>ST4</AUTOGUIDER_PORT>
    <HANDPAD WIRELESS="true" POINTING="false"/>
  </CONTROLLER>
  <MOTOR AXIS="BOTH">
    <VENDOR></VENDOR>
    <MODEL></MODEL>
    <TYPE>Brushless DC Servo</TYPE>
    <GEAR_RATIO>10</GEAR_RATIO> <!-- Always :1 -->
    <RESOLUTION UNITS="STEPS/REV">8000</RESOLUTION>
  </MOTOR>
  <ENCODER AXIS="BOTH">
    <VENDOR>Renishaw</VENDOR>
    <MODEL></MODEL>
    <RESOLUTION UNITS="BITS">26</RESOLUTION>
  </ENCODER>
  <WORM AXIS="RA"/>
</MOUNT>

Listing 8. AISML MOUNT with extensions.
Listing 9. AISML FOCUSER with extensions.

```xml
<?xml version="1.0" encoding="utf-8"?>
<FOCUSER ID="TDSF001">
  <LOCATION>Secondary</LOCATION> <!-- Usually Focal Plane or Secondary -->
  <!-- Below are user additions -->
  <VENDOR>Robofocus</VENDOR>
  <MODEL>Van Slyke</MODEL>
  <SERIAL>N/A</SERIAL>
  <INTERFACE ASCOM="true">Serial</INTERFACE>
  <STEPSIZE UNITS="microns">2.54</STEPSIZE> <!-- 10,000 steps/inch -->
  <TEMPCOMP AVAIL="false" />
  <ABSOLUTE AVAIL="true" />
</FOCUSER>
```

Listing 10. AISML FILTER_WHEEL with extensions.

```xml
<?xml version="1.0" encoding="utf-8"?>
<FILTER_WHEEL ID="TDSFW001">
  <VENDOR>Apogee</VENDOR>
  <MODEL>AFW50-9R-DEEP</MODEL>
  <INTERFACE ASCOM="true">USB 2.0</INTERFACE>
  <DIAMETER UNITS="mm">50</DIAMETER>
  <SHAPE>Circular</SHAPE>
  <SLOTS>9</SLOTS>
  <!-- RANGE set between 50% cutoff -->
  <FILTER ID="01">
    <DESCRIPTION>Sloan r prime</DESCRIPTION>
    <NAME>SR</NAME>
    <MIDPOINT UNITS="micron">0.629</MIDPOINT>
    <RANGE UNITS="micron">0.065</RANGE>
  </FILTER>
  <FILTER ID="09">
    <DESCRIPTION>Clear</DESCRIPTION>
    <NAME>Clear</NAME>
  </FILTER>
</FILTER_WHEEL>
```

Listing 11. Simple AISML AIS.

```xml
<?xml version="1.0" encoding="utf-8"?>
<AIS xmlns:xi="http://www.w3.org/2001/XInclude" ID="AISi001">
  <AIS_OTA> <xi:include href="t_TDS24.xml"/> </AIS_OTA>
  <AIS_CAMERA> <xi:include href="c_ALTAU6.xml"/> </AIS_CAMERA>
</AIS>
```
Listing 12. AISML AIS in Listing 11 with a guide OTA and CAMERA.

```xml
<?xml version="1.0" encoding="utf-8"?>
<AIS xmlns:xi="http://www.w3.org/2001/XInclude" ID="AISi001">
  <AIS_OTA TYPE="Primary">
    <xi:include href="o_TDS24.xml"/>
  </AIS_OTA>
  <AIS_OTA TYPE="Guide">
    <xi:include href="o_TDS05.xml"/>
  </AIS_OTA>
  <AIS_CAMERA TYPE="Primary">
    <xi:include href="c_ALTAU6.xml"/>
  </AIS_CAMERA>
  <AIS_CAMERA TYPE="Guide">
    <xi:include href="c_ATIK16IC.xml"/>
  </AIS_CAMERA>
</AIS>
```
Abstract

For the past two years, The University of Colorado, in collaboration with Las Cumbres Observatory Global Telescope Network (LCOGTN) has been taking Sloan’s r’ and i’ images of approximately 200 galaxies during each new moon period to provide ground data in support of the approximately 1100 hours of warm Spitzer time awarded to Dr. Mansi Kasliwal’s Caltech SPIRITS program. Currently there are over 6,000 images in this archive. Small telescope scientists routinely image the same fields, building similar archives brimming with science potential. This paper reports the technique to develop serendipitous observations of dwarf field stars. Answers to questions surrounding the dwarf’s early life in proximity to non-hierarchal multiple star groups, about how dwarfs not only survive but are so numerous are well within the capabilities of small telescope scientists. The role of the small telescope scientist is of vital importance in these (re)discovery, confirmation, monitoring and reporting tasks.

1. Introduction

This paper uses 30 exposures of NGC 300 (00:54:53.479 -37:41:03.77, 19.6 Mpc, 7.18 x 10^9 L_\odot) as the example for describing a dwarf candidate identification technique.

The profusion of lone dwarfs in galaxies, especially dwarfs with planetary systems, flies in the face of the strong rejection of the idea that stars form in isolation. Dwarfs are numerous cool objects having a thermal peak in the NIR and show a slope, indicated by a large color excess in r’ - i’ bands (Hawley et al. 2002). The sign and degree of this slope is an indicator of where the thermal peak occurs along the Wein spectral side (blue) of black-body curve and serves as a basis for a quick and efficient approach to identify candidate dwarf stars. Observations simply consist of stacked exposures with sufficient signal-to-noise (S/N) in r’ and i’ to achieve a 3-sigma observation.

Approximately 6000 images in r’ and i’ were obtained through the generous support from the Las Cumbres Observatory Global Telescope Network (LCOGTN) in support of the University Colorado (CU) program. CU monitored these images for visual counterparts to the Caltech SPIRITS program of Kasliwal et al. (2014) observations of transient intermediate luminosity events in the ~200 star forming galaxies under consideration. Of the program targets, 182 have declination north of -20. This makes them easily accessible by northern hemisphere observers.

The fields of these images are a mixed bag. The size of the target galaxy can be large, with galaxy texture interfering with the photometry. Some targets are near the galactic plane, others in areas of sparser foreground stars. The images display visible defects that include clear lines where the gain of the chip’s individual channels were not completely managed. Accurate pointing was an issue early in the program with wide margins where the images did not overlap. High clouds, ice, wind, glare and other culprits plagued observations.

The target fields yield rich collections of stars where the color index of r’–i’ ~ 3 (see Hawley et al. 2002) is a strong indicator of a cool star. The range of this magnitude difference yields a low-resolution spectral classification of stellar type. Dwarf variability is on the order of minutes to days with some stars indicating episodes of activity (Bell et al., 2012 references therein). The cadence of the CU observations are not suited to a refined variability study, only to finding candidates that show promise. The overall magnitudes offer targets worthy of follow-up observations with higher cadence. The small telescope science community, including the “pretty picture” crowd are encouraged to collect and analyze these field stars and to share their images with collaborators.

Cooler stars are better measured in the NIR. IR sensors remain out of reach of all but the best funded small telescope scientists at this point. The Sloan’s z’ filter falls into a very low QE range of popular silicon sensors. Working the r’ and i’ bands are still well
within the Wein regime of the blackbody curve for these cool targets, where the blue side of the thermal peak actually extends well into the J, H and K IR bands.

2. Data Analysis

The sample of images were obtained with LCOGTN 1-meter telescopes which are located around the world. They are scheduled with a queue allocation program. These data were fully embargoed, reduced and deposited in the Infrared Processing and Analysis Center (IPAC) repository for the CU program. Two main cameras were used together with two binning protocols. Python scripts together with MySQL, and later PostgreSQL, databases were used to pull header data and compare with the Spitzer transients to produce optical upper-bounds for these events in the visual bands. Roughly 130 header values are embedded in each image. The main tools applied to these data for this paper include Astrometry.net Lang et al. (2010); Source Extractor (hereafter SEXtractor) Bertin (2015), Holwerda (2015); Python scripts together with Astropy Astropy Collaboration et al. (2013); running in the Ureka environment Hirst et al. (2014); SAOImage (ds9) Vanhilst (1990); and the PostgreSQL database. A handful of useful PostgreSQL functions were developed.

2.1 Analysis

The telescope produces raw images (figure 1-step 1), and are processed by an auditable pipeline process (step 2) to make the analysis images (step 3). We pick up the LCOGTN images at step-3. As the supplied coordinates were coarse and not useful, Astrometry.net is used to refine astrometry (steps 4). These files are then processed by SEXtractor (step 5) to produce a catalog file reporting x, y, RA, Dec, FWHM, ellipticity, MAG_AUTO flux, and background estimates for all targets found. Scripts (step 6) select star like targets where FWHM matches the night and ellipticity is 'round' within the errors of the PSF for the images. Additional scripts (step 6) produces the database tables from the SEXtractor and UCAC4/APASS Zacharias & Gaume (2011); Henden et al. (2009) catalog files. The database engine (step 7) uses stored views and procedures to process and produce the report (step 8). The functions are developed to select related targets, match them to the UCAC4 catalog and calculate zeropoints. The refined star data are stored and selected for the report. Observations across different epochs are used to spot variability. A later observing campaign, with a high cadence, can fully report the nature of the variability.

The galaxies in the study are relatively close (5-20 Mpc) and essentially fill their image fields. Care was exercised w.r.t. accepting candidates due to the background "texture" from the edges of the galaxies. The SEXtractor MAG_AUTO aperture technique was used. For a general overview of issues related to doing aperture photometry see Sonnett et al. (2013).

The software used for all steps are free of charge.

As an example: the test image, "lsc1m004-fl04-20141021-0140-e90.fits", reports a FITS header value "L1FWHM" of 2.618 (an average of around 4.6 pixels) and a mean ellipticity "L1ELLIP" of 0.18154. SEXtractor reports "L1FWHM" of 2.7190 and "L1ELLIP" 0.135333 directly. A typical SEXtractor summary is, "Measuring from: "NGC0300", (M+D) Background: 152.288, RMS: 15.7834, Threshold: 39.4586 for 948 detected objects while extracting 886 objects". The 4x4 pixel array yields a very good PSF for stars.

An internal database of dwarf star classification by spectral type was created from information gleaned from the web Reid, N. (2007) and summarized in table 1. Note the overlaps need additional filters to more fully resolve. Since z' band images are an issue, the use of g' filter would make an excellent addition to the mix.

![Figure 1: High-level flowchart of data collection, information retrieval and processing. The raw images are processed by conventional means and refined astrometry added by Astrometry.net and SEXtractor. The PostgreSQL database manages data and rules for producing products. The steps are discussed in section 2.1 of the text. (Image produced with Gliffy.)](image)
An average of 9 candidate dwarfs meeting the criteria of the r'-i' delta-magnitude for the sample of stars around NGC 300 were found for each month's data. The complete list of candidates is in table 2 candidates for further study.

3. Results

The results are to within 0.2 magnitude with the main errors arising from initial photometric reduction of the images and from the variability in zeropoint calculations across the catalog of reference stars for this field. A range of 22.09-23.06 spanned the i' images, and a range of 21.79-25.25 spanned the r' images. No zeropoint color corrections were reported by the observatories for these images.

The dwarf stars are highly variable and the low cadence of observations are only useful to spot stars that may exhibit magnitude changes beyond 0.2 magnitude. Obviously, low cadence of the observations might completely miss stars which are highly variable, just quiet during the time the images were taken.

The following query produced Table 1 showing candidates matching a broad rule of color between 1.0 and 3.63 mag, using a rough interpretation of dwarf classification values from Hawley et. al.

```
SELECT
r2s(ora),d2s(odec),delta,ip,rp,mjd_obs
FROM (SELECT DISTINCT ON (a.ora)
a.ora, a.odec,
(a.mag - b.mag)::numeric(4,2) AS "delta",
(a.mag)::numeric(4,2) AS "ip",
(b.mag)::numeric(4,2) AS "rp",
a.mjd_obs from idata as a
JOIN rdata b
ON vnearby(a.ora, a.odec, 5, b.ora, b.odec)
WHERE abs(b.mjd_obs - a.mjd_obs) > 0.8 AND
a.mag - b.mag between 1.0 and 3.63 AND
a.mag != 'NaN'::float and
b.mag != 'NaN'::float
ORDER BY ora,odec,a.mjd_obs ) xx ;
```

These results overlay a deep composite i' image in Figure 2. The north-east corner appears in Figure 3.

![Figure 2: NGC 300 Sloan's i-band. This is a deep combined image of all i' images showing the locations of potential candidates. Those near the galaxy structure need special attention due to underlying influences of NGC 300's HII and other potential very red underlying sources. The color r - i value is shown.](image)

![Figure 3: NGC 300 Sloan's i-band zoomed into the NE corner of the image. Note two or more magnitude variations were detected by the process and are shown by 'overlapping' circles. The color r - i value is shown.](image)
4. Discussion

Peak temperatures for dwarf stars occur in the mid infrared – beyond the budgets of small telescope research. By working the Sloans $r'$ and $i'$ filters it is possible to gleam dwarf observations from unrelated non-dwarf programs. Statements about serendipitous discoveries about small, cool and relatively dim stars are dependent on the S/N of the original research images.

The astrometry of these stars is of interest to the community as well. No stars form in isolation. A large sub-population of dwarfs form together with target O and B stars in early birth clusters. The brighter candidates are therefore closer to Earth and provide a great opportunity for proper motion studies. Astrometry for the stars reported from the 2MASS and Sloans studies are worth the investigation. These data are available on-line. In figure 1 step-6 additional tab-separated-variable files can be input into the database; step-7's views can be extended to handle astrometric reporting.

Dwarf stars march to their own drummers. To separate those exhibiting greater variability provides sub-lists to support different observing goals. While stars with high variability are of little interests to astrobiologists, they offer a great laboratory for stellar physics. Another small telescope dwarf project is described by López-Morales & Clemens (2004).

5. Conclusion

Vast numbers of images exist in public databases holding a wealth of un-refined and un-reported data. The small telescope scientist is encouraged to utilize these data and to report their approaches and findings for use by educators and other small telescope researchers. The results from this project suggests opportunities exist to follow up with galaxies to contribute a higher cadence program of collaborative observations to major research projects. The image archives are a great starting point for the beginning small telescope scientist. Here one can develop and refine techniques while acquiring a great skills base and saving funds to acquire equipment. Observatory Global Telescope Network (LCOGT), the excellent tool suites provided by STScI and the greater astronomy software community. The LCOGTN IPAC archive is courtesy of NASA/JPL-Caltech.

6. Acknowledgements

The author would like to thank the University of Colorado for the collaboration opportunity with the SPIRITS program, to acknowledge the generous allocation of queue time from the Las Cumbres Observatory Global Telescope Network (LCOGT), the excellent tool suites provided by STScI and the greater astronomy software community. The LCOGTN IPAC archive is courtesy of NASA/JPL-Caltech.
7. References


Early Images of Sodium in the Tail of Comet Hale-Bopp

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Abstract

Astronomers announced that sodium was discovered in the extended tail of Comet Hale-Bopp on 16 April 1997. A literature search uncovered three reports of sodium near the coma in February. Two additional references reported slit spectra of sodium in the extended tail on 09 March and 13 April, but neither reported the orientation of a sodium tail. This paper presents results of film spectra taken of Comet Hale-Bopp by the author with a spectrograph that he had designed and built. Images of ion tails were recorded between 09 March 1997 and 05 May 1997. Seven showed the extended sodium tail. The spectrum of 09 March is presented alongside a corresponding star chart. Several spectra show two streamers of ion tails and a singular sodium tail. The orientations of these tails were measured, tabulated, and shown graphically, along with the antisolar directions.

1. Introduction

In late 1986 Comet Hale-Bopp (C/1995 01) appeared to public view, became the brightest comet in decades, and maintained easy visibility throughout half of 1997. After observing Comet Hale-Bopp with a 50-mm telescope on 16 April 1997, three astronomers; Cremonese, G., Fitzsimmons, A., and Pollacco, D. (Sky & Telescope, 1997; Astronomy, 1997) announced the discovery of two long sodium tails, one wide and diffuse, and the other sharp. The author, having taken film spectra of the comet with a 71 mm objective lens, beginning on 09 March 1997, examined the spectra. Many of the images showed the sodium tail as early as 09 March 1997, 38 days prior to the announced discovery of sodium in the extended tail. Some of these images were made before the perihelion, which occurred on 27 March. This paper reports the orientations of the ion and sodium tails and the antisolar direction for various dates.

2. Reports of Sodium in Tail before 17 April 1997

Furusho, R. et al. monitored spectra from Sep 1996 to May 1997 within a 7.4” by 7.4” aperture, which covered only the cometary central condensation. Sodium D-line emission was presumably detected in early Feb 1997 and for six additional days before perihelion passage. (Furusho, 2005)

M. Sakamoto et al. found neutral sodium emission images of the comet since 20 February 1997, in size of 500,000 km by 500,000 km. (Sakamoto, 1997)

T. Kawabata and K. Ayani, Bisei Astronomical Observatory, reported emission lines of Na D1 and D2 on Feb. 26.9 within 7” of the Hale-Bopp’s nucleus. (Kawabata, 1997)

G. Avila et al. reported that six slit spectra of the Comet Hale-Bopp tail were recorded by the European Southern Observatory and Max-Planck Institute for Extraterrestrial Physics near Munich during the period 9 March to 24 April 1997. The sodium line is present on 9 March, 21 April, and 24 April, but not present on 1, 7, and 8 April. The presence in the spectrum of 09 March may be due to the background light of close by cities, according to the writers. No orientation of the sodium tail was reported. (Avila, 1997)

R. Heyd et al. obtained a spectrum of the coma and tail on 13 April. The sodium emission line appeared in both parts of the spectrum, but no orientation of the tail was obtained. (Heyd, 1997)

G. Cremonese, Astronomical Observatory, Padua, reported an extensive neutral Na tail structure associated with Comet Hale-Bopp with a wide-field 50 mm instrument through a Na D filter on Apr 16.88 UT at the Observatorio del Roque do los Muchachos, La Palma, 3 deg north of the projected antisolar direction. (a broad ionized Na tail and a narrow neutral Na tail) G. Cremonese et al. observed the tail was between 6 and 7 degrees long, corresponding to a length of about 50 million km. (Cremonese, 1997)

J. K. Wilson et al. observed the width of the Na lines in the tail on 17 March and 20 March 1997, and determined the positions of them in relation to the dust tail and the ion tail. (Wilson, 1998).
3. The Spectrograph and its Operation

The author designed and built a versatile hand-portable spectrograph that is capable of recording reference marks alongside spectra taken in the slitless mode. This self-contained spectrograph is useful for obtaining spectra of many light sources under various field conditions. The transmission diffraction grating has 600 grooves per millimeter and is blazed for 4800 Å in the first order. The geometry and mathematics are described elsewhere (Buchanan, 2016).

The first beam of light from the distant target object strikes the diffraction grating and passes through a camera lens, which focuses the dispersed light onto the film and creates the spectrum. Three optical paths for zeroth order are provided by mirrors and lenses to provide the reference marks along the slitless spectrum. The first reference mark is placed near the Fraunhofer-B line of oxygen, 6870 Å (Weast, 1967). The second reference mark is placed near the hydrogen-zeta line, 3889 Å (Ferrence, 1959; Harrison, 1969). The third reference mark allows the making small corrections in the positions of the other two reference marks. Positional error in the 6870 Å mark never exceeds 7 Å. Positional error in the 3889 Å mark never exceeds 3.5 Å. The reference marks apply to the first order, when the aim toward the light source is less than two degrees in error. The 6870 Å line can apply as 3435 Å in the second order. The reference marks can be used in a similar manner to the lines of a comparison spectrum for a slit spectrograph.

The objective lens has a focal length of 135 mm and an aperture of 71 mm. It can record the entire visible first order in one exposure. The camera can be rotated about the vertical axis of the grating and fastened at any of seven positions, making possible the imaging of two parts of the second order and five parts of the third order. Appropriate filter(s) for each position will minimize the overlap of orders. The most ideal resolution is approximately 3.0 Å in the first order, 1.5 Å in the second order, and 0.9 Å in the third order. Only the first and second orders were obtained of Comet Hale-Bopp.

4. Effects of Sky Glow on Spectra

For slitless operation, a shield containing a circular hole reduces the effects of sky glow. For camera apertures of f/1.9 and f/2.8, the appropriate shield is inserted at the front of the spectrograph structure, about 0.5 meter in front of the diffraction grating. Each shield has a hole diameter about 5 millimeters greater than the corresponding camera aperture, so that all wavelengths from the target object will cover the diffraction grating.

The sky glow is strongest in the orange, due to the prevalence of high-pressure sodium vapor lamps. A focused spectrum of this lamp type contains the sodium doublet as a pair of dark absorption lines. However, in a slitless spectrograph, the sky glow is unfocused, resulting in a diffuse fogged elliptical area on the film.

5. Obtaining and Measuring Spectra

5.1 Description of Equipment and Operations

The observing site was located about 32 miles west of downtown Atlanta, Georgia. The low northeastern sky lay within the sky glow of the northern Atlanta metropolitan area. The spectrograph was fastened to its German equatorial mount in such manner that the spectrum spread only in the plane through the axis of the equatorial mount. An attached four-power spotting telescope with cross hairs was used to align the spectrograph toward the target. During most exposures the battery-powered motor drive was operated at 100% sidereal rate. However, in order to widen the spectrum on a few images, the drive motor was stopped for one second at constant intervals on the exposures noted on Table 1. Except for the calibration exposures of AlpCMa, also known as Sirius, and of Comet Hale-Bopp, the three beams causing reference marks were blocked to eliminate possible interference with the spectral images, since the spectral reference mark of 3889 Å is close to the 3883 Å line of the comet. Table 1 contains details of each spectrogram successfully exposed.

Under normal operation, the spectrum spreads along a right ascension arc, passing through a celestial pole. This arrangement showed the comet's tail well when the tail was pointed westward or eastward. However, from late March to early April, the tail was pointed generally southward. In order to spread the spectrum east to west, a star was selected which lay along an arc perpendicular to the right ascension arc passing through the comet. The mount was oriented to make the spectrum spread along the plane of this arc. During exposure, the declination knob was turned manually at the rate necessary to maintain the coma on the cross hairs. Exposure times in these cases were limited to 8 minutes in order to avoid significant deviation from the path of the comet along the curved declination line.

Kodak technical panchromatic film was used for all exposures. Film rolls #’s E-145, E-146, and E-
were hyper-sensitized. Roll # E-146 was not hyper-sensitized. The films were scanned by use of a Pacific Image PrimeFilmXA scanner. In some cases a slide duplicator was used with shading of lightly exposed areas of the film so that the heavily exposed areas would show more clearly.

5.2 Details of Useful Spectra

A list of dates, times, and other details are described in Table 1.

<table>
<thead>
<tr>
<th>Photo ID</th>
<th>Film Type</th>
<th>Subject Name</th>
<th>Date Taken</th>
<th>Time Taken</th>
<th>Shutter Speed</th>
<th>Aperture</th>
<th>Driven % Sidereal</th>
<th>Shield Hole Dia.</th>
<th>Filter</th>
<th>Order</th>
<th>Camera Position</th>
<th>Spectrum Direction</th>
</tr>
</thead>
<tbody>
<tr>
<td>E-145-01*</td>
<td>Kodak TP Hypo</td>
<td>Comet H-B</td>
<td>01-Mar-97</td>
<td>09:55</td>
<td>3m</td>
<td>1/4</td>
<td>100</td>
<td>No shield</td>
<td>1</td>
<td>N-S. violet-red</td>
<td></td>
<td></td>
</tr>
<tr>
<td>E-145-02*</td>
<td>Kodak TP Hypo</td>
<td>Comet H-B</td>
<td>08-Mar-97</td>
<td>10:05</td>
<td>8m</td>
<td>1/4</td>
<td>100</td>
<td>75mm</td>
<td>1</td>
<td>N-S. violet-red</td>
<td></td>
<td></td>
</tr>
<tr>
<td>E-145-03</td>
<td>Kodak TP Hypo</td>
<td>Comet H-B</td>
<td>06-Mar-97</td>
<td>10:18</td>
<td>3m</td>
<td>1/28</td>
<td>100</td>
<td>54mm</td>
<td>1</td>
<td>N-S. violet-red</td>
<td></td>
<td></td>
</tr>
<tr>
<td>E-145-10</td>
<td>Kodak TP Hypo</td>
<td>Comet H-B</td>
<td>16-Mar-97</td>
<td>09:40</td>
<td>3m</td>
<td>1/28</td>
<td>100</td>
<td>54mm</td>
<td>1</td>
<td>N-S. violet-red</td>
<td></td>
<td></td>
</tr>
<tr>
<td>E-145-11</td>
<td>Kodak TP Hypo</td>
<td>Comet H-B</td>
<td>16-Mar-97</td>
<td>09:42</td>
<td>4m</td>
<td>1/28</td>
<td>100</td>
<td>54mm</td>
<td>1</td>
<td>N-S. violet-red</td>
<td></td>
<td></td>
</tr>
<tr>
<td>E-145-12</td>
<td>Kodak TP Hypo</td>
<td>Comet H-B</td>
<td>16-Mar-97</td>
<td>09:47</td>
<td>8m</td>
<td>1/28</td>
<td>100</td>
<td>54mm</td>
<td>1</td>
<td>N-S. violet-red</td>
<td></td>
<td></td>
</tr>
<tr>
<td>E-145-13</td>
<td>Kodak TP Hypo</td>
<td>Comet H-B</td>
<td>16-Mar-97</td>
<td>09:58</td>
<td>7m</td>
<td>1/28</td>
<td>100</td>
<td>54mm</td>
<td>1</td>
<td>N-S. violet-green</td>
<td></td>
<td></td>
</tr>
<tr>
<td>E-145-15</td>
<td>Kodak TP Hypo</td>
<td>Comet H-B</td>
<td>16-Mar-97</td>
<td>10:10</td>
<td>15m</td>
<td>1/28</td>
<td>100</td>
<td>54mm</td>
<td>1</td>
<td>N-S. violet-green</td>
<td></td>
<td></td>
</tr>
<tr>
<td>E-145-22</td>
<td>Kodak TP Hypo</td>
<td>Comet H-B</td>
<td>30-Mar-97</td>
<td>01:49</td>
<td>8m</td>
<td>1/28</td>
<td>100</td>
<td>54mm</td>
<td>1</td>
<td>N-S. violet-green</td>
<td></td>
<td></td>
</tr>
<tr>
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<td>Kodak TP Hypo</td>
<td>Comet H-B</td>
<td>30-Mar-97</td>
<td>01:49</td>
<td>8m</td>
<td>1/28</td>
<td>100</td>
<td>54mm</td>
<td>1</td>
<td>N-S. violet-red</td>
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<td></td>
</tr>
<tr>
<td>E-145-32</td>
<td>Kodak TP Hypo</td>
<td>Comet H-B</td>
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Table 1: Useful Spectra of Comet Hale-Bopp

Figure 1. Spectrum of Comet Hale-Bopp on 09 Mar 1997, with unreadable spectral streaks of stars that are shown on chart at right. Map from Uranometria 2000.0 by Tirion, Rappaport and Lovi, Copy right © 1987, 1988 Willmann-Bell, Inc. Used with permission.

Figure 3. Juxtaposed spectra of Comet Hale-Bopp with its overlying ions and AlpCMa, with reference marks aligned between the two. A slit spectrum of the sun underlies the other spectra.
All spectra were taken slitless. All spectra except the E146 series were taken with hypersensitized Kodak technical pan film. The spectra in the E146 series were taken with similar film without hypersensitizing. The intent was to emphasize the spectrum of the coma and not to overexpose the spectrum AlpCMa.

6. Data and Results

6.1 Two Strong Streamers of CO⁺ Ion Tail

Published photographs by Others show that the ion tail displayed two strong streamers on most dates from 16 February to 10 April. The author examined 20 photographs by various people as published in ‘Astronomy’ and ‘Sky & Telescope’ in several monthly issues and elsewhere (JPL/NASA). The approximate angles between these streamers were read to the nearest degree by use of a protractor on The angle opened from 8 degrees on 16 February to 13 degrees on 30 March, and then closed to 11 degrees on 10 April. The edges of the strong streamers were indistinct, making the measurements approximate. On a few dates the identity of separate streams was indistinct.

6.2 Identifying Sodium in the Tail

The positions of the spectral lines on the various scanned images were measured in Adobe PhotoShop CS6. The streamers of the comet tail were measured eastward (counter-clockwise) and westward (clockwise) from north. Sodium was identified in Comet Hale-Bopp's tail by its relative position along the spectrum as compared with known spectra containing sodium. Due to low contrast of the tail images and to the indistinct boundaries, these measurements are not precise, but are assumed to represent the actual orientations within about 2 degrees.

Figure 1, taken on 09 March 1997, shows the sodium output toward the upper left, with a short tail image. All spectra in this paper cover the range approximately from 3850 Å to 6950 Å. The Sodium appears at 5893 Å near the 2/3 point along the spectrum. The thin horizontal streaks on the images are unreadable spectra of stars near the comet, the locations of which match stars on the star map at right. The spectrum of the coma was overexposed, but some of the ion tails were evident.

Three spectra were taken on 16 March with lighter exposure. The spectrum of the coma could be read, including the sodium thereon, but exposure was insufficient for the ion tails and sodium tail to show.

On 30 March the tails were pointed close to the North Pole, so that the normal operation of orienting the spectrograph N-S or S-N would not show them well. Instead, the spectrograph was advantageously oriented at right angles to the usual.

Figure 2. A Spectrum of Comet Hale-Bopp on 30 March 1997, showing ion tails and sodium tail

Figure 2 shows that the tails had become brighter and longer. Also, the ion tails had two streamers several degrees apart, while the sodium tail was singular.

Figure 3 shows a spectrum of Comet Hale-Bopp taken on 08 April by a driven rate slower than sidereal. A reference mark is shown at each end so it could be compared with the underlying spectrum of AlpCMa taken the same evening. The left and right reference marks were calculated to be at wavelengths of 3872 Å and 6916 Å. To further identify sodium, a slit spectrum of the sun, taken at another date lies just below the spectrum of AlpCMa. Immediately above the spectrum of the coma, spectra of gas output are evident. The sodium output originates at the spectrum of sodium in the coma and extends toward the upper right.

6.3 Positions of the Sodium Tail and the CO⁺ Ion Tail

The positions of the comet were obtained from references (Ephemeris of Comet Hale-Bopp, Yeomans, 1997). The positions of the sun are listed in an almanac (The Astronomical Almanac for the Year 1997) The antisolar directions from the comet were computed by spherical trigonometry.
Table 2. Orientations of Ions and Sodium Tails

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Table 2. Orientations of Ions and Sodium Tails

All ion streams and the sodium tail stayed close to the antisolar direction. The sodium tail appears to stay closer to the west ion stream, but appears to be independent of both east and west ion streams.

The author previously obtained spectra of Comet Halley and contributed them to International Halley Watch (Edberg 1996). The author examined his spectra of Comet Halley in 1986, Comet Bradfield in 1987, Comet Levy in 1990, and Comet Hyakutake in 1995. No sodium tails could be detected in any of these spectra.

7. Discussion and Conclusions

1. The ion tail generally had two streams. The eastern and western limbs were separated by about 11 to 14 widening to about 23 degrees on 01 May.

2. The image taken of the sodium tail on 09 March 1997 pre-existed the announced discovery of sodium in the tail of Comet Hale-Bopp on 16 April 1997. The faintness of the images taken on 09 March was probably due to the sky glow in the northeast where the comet lay before dawn. Later images that displayed the sodium tail more distinctly were taken after dusk of the comet in the northwest where less sky glow existed.

3. Although sodium has been detected near the nucleus in comets many times before, but this is the first time that a tail of Na outside the nucleus has been discovered (Astronomy, 1997), (Sky & Telescope, 1997).

8. Acknowledgments

Marsden, B. for providing accurate orbital elements for the comet.
9. References and Bibliography

Astronomy, An Unexpected Tail for Hale-Bopp, August 1997, 22.


Ephemeris of Comet Hale-Bopp
http://www2.jpl.nasa.gov/comet/ephemjpl8.html


Sky & Telescope, Hale-Bopp’s Sodium Tail, July 1997, 16


1. Introduction

A clear 5.2-year photometric cycle ($\Delta V \approx 0.14$ mag) in the object PG1302-102 was recognized by Graham et al (2015), based on data from several surveys using unfiltered CCD photometry (primarily the CRTS data). Their preferred hypothesis is that this signal represents the orbital period of a binary black hole. In a personal communication, Dr. Graham indicated that multi-band monitoring of this object would be interesting, because there is no good data about how its color might change during the brightness cycle (if at all). He laid out pretty challenging goals for accuracy: $\pm 0.02$ mag desired, and $\pm 0.05$ mag required.

Subsequently, D’Orazio et al (2015) suggested that the mechanism for the brightness cycle could be “relativistic boost”: the large Doppler shift generated by the orbital speed of the two black holes makes the spectral energy distribution shift red-ward and blue-ward during the orbital cycle. In this scenario, the brightness change should be greater in the UV than it is in the visible band, and they do find evidence for this in the available near- and far-UV data.

Jun et al (2015) examined IR data. Although the data is sparse, the amplitude in the IR seems to be substantially lower than it is in the visible band.

The results from D’Orazio and Jun seem to encourage multiband photometry monitoring, with the idea that the source’s fluctuation in B-band may be larger than it is in V- or R-band. That is the goal of my long-term photometry project, although it remains to be seen whether my equipment is up to the challenge.

2. Imaging Plan

The field of PG1302-102, along with the comp stars identified in the AAVSO photometry table (derived from APASS) is shown in Figure 1. The comp star identifiers (A through T) were assigned for convenience. I used star N as the comparison star for differential photometry, with star E and star B used as check stars.

PG1302-102 is a challenging target from my observatory: it culminates at elevation angle of 45 degrees (AM=1.4); it is in the south, which is where both light pollution and the risk of marine fog are high; at Vmag $\approx 15$ it is pretty faint.

Initial experiments at my Altimira Observatory using 5-minute exposures, yielded SNR on the target image of SNR(b-filter) $\approx 25-65$, SNR(v-filter) $\approx 50-90$ and SNR(r-filter) $\approx 45-75$, depending on the clarity of the night and the air mass of observation. Since $\sigma_{\text{phot}}= 1.0857/\text{SNR}$, it appears that the project goals are (barely) achievable; and encourages using exposures of about 30 minutes to improve effective SNR.

My setup can’t reliably take an exposure longer than 5 minutes (with autoguiding), so my strategy is to take 5X images at 5 minutes each, in B and V (on some nights) or V and R (on other nights), and to devote the entire night to this target (thereby
collecting an hour or two in each filter on each night). The images are reduced with bias, dark and flat frames, then combined by “align-and-sum” using Maxim-DL. Differential photometry is done with MPO Canopus.

Part of the challenge to finding the instrumental color index is that the B and V images are not taken at the same time – indeed, with my more-or-less standard imaging sequence (5 B-band images, then 5 V-band images), they are roughly a half-hour apart and all sorts of atmospheric changes can happen during that time (air-mass difference, obviously, but temporal transparency changes are probably a greater danger). By taking several image sequences, the nightly data is V-B-V-B-V-B, which permits a linear interpolation to create “synthetic” instrumental magnitudes simultaneously in both filters. The concept of this is illustrated in Figure 2. For each “B” instrumental magnitude, I use the preceding “V” and the following “V”, and linear interpolation to estimate the “V” instrumental magnitude that would have been observed simultaneous with the B image. In general, this seems to work well; by examining the data it usually become clear if something went awry -- e.g. larger-than-normal temporal changes in the instrumental magnitude between two sum-images.

Within the range of validity – i.e., the range of star colors used to create the transform – the two regression lines are essentially identical. But if you extrapolate to a much bluer or redder instrumental (b-v) color, then the calculated standard [B-V] depends strongly on which of the two “essentially identical” regression lines you use. This isn’t a great revelation, that extrapolation is a very dangerous activity, but its worth remembering.

The need for a range of colors extending into the blue limits the choice to Landolt fields to use. Happily, it Landolt field PG1323-086 is both near the target (only 5 degrees away) and contains two very blue stars (PG1323-086A has color index [B-V]= 0.393, and PG1323-086 itself has an extraordinarily blue color index of [B-V]= -0.14). However, PG1323-086 is (like PG1302-102) a quasar; and some quasars display random brightness fluctuations on a variety of time scales; so it leaves you with a nagging uncertainty about whether it’s a very good standard star. Such nagging uncertainties aren’t necessarily signs of paranoia: Clem and Landolt (2010) reported that PG 1323−086A is an eclipsing binary with a lightcurve amplitude of ≈ 0.25 magnitude. So two of the 4 standard stars in this field are tarnished.

The solution that I chose was the field of SA29, which is overhead at about the time that PG1302-102 is rising. The bluest star in SA29 is SA29-167, at [B-V]= 0.398. The other stars in this field range from [B-V]=0.522 to [B-V]= 1.071, so it offers a nice wide
span of color indices, and demands only a modest extrapolation in color index.

3. Differential Photometry

The analysis approach is routine differential photometry, based on known color index and magnitude of the comp star (see Appendix for the equations used).

I use either the [B-V] or the [V-R] color index to determine the target V-magnitude, depending on which filters I used during the night.

The differential photometry results from MPO Canopus are exported as a text file, and the analysis and graphing is done in an Excel spreadsheet.

4. Calibrating the Comparison Stars

I began with the idea of calibrating the comp stars by doing “all sky” photometry from my observatory. The Landolt fields SA-104 and PG-1323 are conveniently close to the target, so it seemed like a straightforward task. The pitfall was that the target (and these standard fields) never rise very high above my southern horizon (the target culminates at elevation angle of about 45 degrees, equivalent to air mass = 1.5), and culmination azimuth happens to be right where the local marine layer collects and migrates toward my neighborhood. So, a genuinely clear and stable atmosphere in this direction is relatively rare, especially in spring. I’ll eventually have a good night for this experiment, but it hasn’t happened yet.

For now, I’m using the APASS magnitudes of comp stars as an interim measure.

John Hoot has offered a few hours of his allocated time on a remote-access telescope at Los Cumbres in Chile. At this location, the target culminates nearly straight overhead, which hopefully will yield higher-quality calibration of the comp stars.

5. Results

For the results shown here, I’ve used Star N as the comp star, with Star E and Star B as check stars. The comp (Star N) is significantly brighter than the target, so it has high SNR and its contribution to photometric uncertainty is negligible; and it is one of the bluest stars in the field, so it’s as close to a color-match with the target that I can get. The check Star E is fainter than the comp star, but significantly brighter than the target, so it should provide a warning of any low-level variability in the comp star. Finally, check Star B is roughly the same V-mag as the target, so it offers a good surrogate to identify any odd effects in the photometry of a very faint object; however, this check star is significantly redder than the target. I haven’t seen any trend related to the color difference (so far).

The differential photometry results for check star E are shown in Figure 4. The star is constant, although the internal consistency (scatter) is worse than desired in the 2015 data (the left half of the chart).

![Figure 4: Relatively bright check star “E” shows no variability and indicates the internal consistency (scatter) of the differential photometry.](image)

A similar plot of the differential photometry of the check star B is shown in Figure 5; again the trend is flat, as desired, but again, the scatter is greater than desired in the 2015 data.

![Figure 5: Check star “B” is roughly the same V-mag as the target, and shows the scatter in differential photometry for a faint target.](image)
To put my results into a longer context, they are plotted in Figure 7 along with the CRTS data that formed the core of Graham’s (2015) data set, and a sinusoid suggested by Graham to guide the eye along the 5.2-year periodic brightness cycle.

6. Photometric Accuracy/Consistency Estimates

As noted above, the 2015 season data has a larger-than-desired scatter. Is this a problem with the data, the expected level of random fluctuation, or an actual fluctuation in the target brightness? The data itself will provide an estimate of the photometric accuracy being achieved. Start with a simple analysis of accuracy. The color index is determined by comparing the target to a comp star, in two instrumental color bands:

The uncertainty in \([B-V]_{tgt}\) is driven by the uncertainties in each of the terms. As a reasonable approximation, I assume that each term is independent (this may not be strictly true, since there might be correlations between the terms, but ...)

Further, I’ll assume that the comp-star color index is strictly known and invariant (so its \(\sigma\) is zero);

similarly I’ll assume that the air mass is known exactly, and that the second-order extinction is constant. With these assumptions, the uncertainty in any determination of the target color index is:

\[
\sigma_{[B-V]_{tgt}} = T_{BV} \left[ \sigma_{[B-V]_{comp}} + \sigma_{[V-R]_{comp}} + \sigma_{[V-R]_{tgt}} + \sigma_{[V-R]_{comp}} \right]
\]

where the symbol indicates “root-sum-square”, also known as “add in quadrature”.

There are two ways to estimate the sigmas:

One way is to estimate the expected photometric uncertainty, based on the signal-to-noise ratio by

\[
\sigma_{phot} = \frac{1.0857}{SNR} \text{ magnitudes}
\]

This is a handy formula because most differential-photometry programs provide an estimate of the SNR on each star image. A simple propagation-of-errors analysis shows that the expected photometric random uncertainty in color index is

\[
\sigma_{[B-V]} \approx 0.05 \text{ and } \sigma_{[V-R]} \approx 0.04.
\]

The expected random uncertainty in standard magnitude is

\[
\sigma_{[M_B]} \approx 0.03, \quad \sigma_{[M_V]} \approx 0.02, \quad \text{and } \sigma_{[M_R]} \approx 0.02.
\]

The second way to assess the photometric accuracy/random uncertainty is to let the measurements themselves telegraph their internal consistency. Suppose that – after combining images
I have $N$ summed-images of the field. Each sum-image gives a determination of $\Delta b = [b_{\text{tgt}} - b_{\text{comp}}]$. If life and the universe were perfect, then each sum-image would yield exactly the same value of $\Delta b$. In real life, each image gives a slightly different value, because of noise, imperfect dark subtraction, etc. So, given $N$ sum-images from a single night, calculate the standard deviation $\sigma_{\Delta b}$ of the $b$-band differential magnitude and $\sigma_{\Delta v}$ of the $v$-band. A reasonable estimate of the photometric accuracy of $[B-V]_{\text{tgt}}$ based on the internal consistency of the differential photometry is

$$
\sigma_{[B-V]_{\text{tgt}}} = T_{BV} \left[ \sigma_{\Delta b} \oplus \sigma_{\Delta v} \right]
$$

As a practical matter, I calculated both of these for each night, and the two estimates of photometric accuracy were comparable.

Note that using the consistency of the differential photometry to telegraph the photometric accuracy has a hidden assumption — namely, that the target does not display any actual brightness changes on time scales shorter than a day or so. That is still an open question (in my mind, for this project, at least); some quasars are known to fluctuate on a variety of time scales, even displaying intra-night variation. So some of the intra-night variation that I’m ascribing to uncertainty might, in fact, be real fluctuation in the target. I’ll look into that if I ever get a larger telescope.

Despite good SNR on each sum-image in each color band, and good intra-night consistency, the scatter in color index and magnitude from night to night is larger than expected from the SNR-derived photometric accuracy. Probably some of the night-to-night scatter is related to inconstant atmospheric conditions. I have a suspicion that another contributor is “flexure” between my piggyback guide scope and the main telescope. The flexure causes the target to slowly migrate, by an arc-minute or so, across the image over the course of a night. If life were perfect, then this wouldn’t matter. But small flat-fielding inconsistency (I’ve checked – residual flatness seems to run 1-2%) and a similar level of dark-subtraction inconsistency (about which see below), random effects amounting to several hundredths of a magnitude (independent of SNR) aren’t unlikely.

In order to virtually eliminate the “flexure” problem, I installed a Telescope Drive Master to my Celestron fork mount. This high-resolution encoder monitors the tracking rate and sends corrections to the autoguider input, so that the ‘scope maintains an exact sidereal rate, despite motor and gear inaccuracy. This just went into service in March, so I don’t have much evidence, but it seems to be improving the nightly consistency of differential photometry (by keeping the target on the same pixels all night, and hence side-stepping the effects of drift modulated by flat- and dark-frame residual imperfections).

### 7. Conclusions

PG1302-102 has faded noticeably between early 2015 and early 2016. This is in line with the extrapolation of Graham et al.’s (2015) lightcurve cycle, which predicts a brightness minimum during 2016. The brightness seems to have faded more in $B$-band than in $V$- or $R$-band; but the scatter in the data points makes the color change barely believable. More data, with improved accuracy/consistency and more rapid cadence will be needed over the next 5 years to map out a complete cycle of this object.

### 8. Acknowledgements

This work made use of the Catalina Sky Survey Data Release 2. The CSS survey is funded by the National Aeronautics and Space Administration under Grant No. NNG05GF22G issued through the Science Mission Directorate Near-Earth Objects Observations Program. The CRTS survey is supported by the U.S. National Science Foundation under grants AST-0909182 and AST-1313422.

This research was made possible through the use of the AAVSO Photometric All-Sky Survey (APASS), funded by the Robert Martin Ayers Sciences Fund.

This research has made use of the SIMBAD database, operated at CDS, Strasbourg, France.

This research has made use of the VizieR catalogue access tool, CDS, Strasbourg, France.

This research has made use of NASA’s Astrophysics Data System.

### 9. References


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Appendix

Differential photometry equations. Upper case indicates standard magnitudes and color indices. Lower case indicates instrumental magnitudes “as observed” (not adjusted for atmospheric extinction).

\[
[B - V]_{tgt} = [B - V]_{comp} + T_{BV} [(b - v)_{tgt} - (b - v)_{comp}] (1 - k'X)
\]

\[
B_{tgt} = B_{comp} + (b_{tgt} - b_{comp}) + T_{BV} [(B - V)_{tgt} - (B - V)_{comp}] - k_b X [(b - v)_{tgt} - (b - v)_{comp}]
\]

\[
V_{tgt} = V_{comp} + (v_{tgt} - v_{comp}) + T_{BV} [(B - V)_{tgt} - (B - V)_{comp}]
\]

\[
[V - R]_{tgt} = [V - R]_{comp} + T_{VR} [(v - r)_{tgt} - (v - r)_{comp}]
\]

\[
V_{tgt} = V_{comp} + (v_{tgt} - v_{comp}) + T_{VR} [(V - R)_{tgt} - (V - R)_{comp}]
\]

\[
R_{tgt} = R_{comp} + (r_{tgt} - r_{comp}) + T_{VR} [(V - R)_{tgt} - (V - R)_{comp}]
\]
Poster Papers
The Desert Fireball Network

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Meteorites tell exceptionally valuable stories, providing insights about the pre-solar stellar evolution of our galactic neighborhood, the formation and evolution of our solar system’s protoplanetary disk, and the accretion and differentiation of planetesimals and protoplanets in the early solar system.

Fireball camera networks can be especially valuable tools in the study of these objects. Networks of cameras featuring coverage that provides multiple images of a given fireball from multiple locations allow for the precise 3D path of a fireball through the atmosphere to be determined. This can facilitate determining probable locations for any meteorites making it to the ground, and enable quick recovery minimizing terrestrial contamination. Such networks can also be used to calculate the orbits of recorded fireballs prior to their encounter with the Earth. The potential for being able to directly associate physical composition of recovered specimens with asteroid families is especially exciting.

Deserts have proven to be particularly good locations for finding meteorites. Lack of plant cover facilitates location of specimens, and the arid environment limits degradation and can slow terrestrial contamination. The Nullarbor desert of Western Australia’s exceptionally dark, clear skies and arid environmental conditions make it an outstanding location for establishing such a camera network. The Desert Fireball Network (DFN) led by Phillip Bland of Curtin University was established to take advantage of this particularly well-suited environment. A prototype network consisting of four film cameras was established in 2007 (Bland et al. 2012). This prototype covered a relatively small area (172,000 km²), but by 2010, the network had enabled the recovery of two meteorites (Bunburra Rockhole and Mason Gully). The team was able to determine the precise pre-atmosphere orbits for the recovered meteorites (Bland et al. 2009; Towner et al. 2011) and determine that the Bunburra Rockhole specimen had a basaltic achondrite composition and an Aten-type near-Earth object orbit (Bland et al. 2009). This success allowed us to upgrade and expand the DFN.

The current DFN consists of 32 camera stations covering ~1.3 million km². The individual stations are fully autonomous systems, capable of operating for 12 months without maintenance and storing all images collected over that period. Each DFN observatory station features a 36MP full-format consumer DSLR camera with a liquid crystal (LC) shutter instead of a mechanical shutter. Each observatory station is run by an embedded computer. The autonomous nature of the station allows it to calibrate its own optics, conditionally conduct observations based on cloud conditions, automatically recognize fireball events, and pre-processes the data prior to uploading it to the project server. The LC shutter within the camera breaks fireball trails into dashes for velocity calculation, after triangulation. The data pipeline includes image processing to determine fireball position at sub-pixel level, triangulation and fireball trajectory modelling, and dark flight trajectory modified by weather and forecast climate modelling (Howie et al 2015).

The Desert Fireball Network is now benefitting from a partnership with NASA’s Solar System Exploration Research Virtual Institute (SSERVI). As a virtual institute, SSERVI funds investigators at a broad range of domestic institutions, bringing them together along with international partners via virtual technology to enable new scientific efforts. The Desert Fireball Network and its SSERVI partners are working to expand the DFN to additional locations throughout the U.S. desert southwest as well as appropriate locations around the world. We are now working to identify potential partners and locations for this expansion.

The Desert Fireball Network includes an integral citizen science program, Fireballs in the Sky (www.fireballsinthesky.com.au). A mobile app (www.fireballsinthesky.com.au/download-app) has been developed for Android and iOS so that the public can record and share their fireball sightings with DFN scientists and participate in real-time research. Fireballs in the Sky has the goal of sharing DFN research and getting students and the general public directly involved in DFN science. Participants are encouraged to contribute and interact with DFN scientists and become members of the DFN Fireballs in the Sky community (Bland et al 2015).

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Abstract
The purpose of this experiment was to map the magnitude and temperature of RR Lyra throughout a complete pulsation cycle to determine if there is a relationship. RR Lyr is an intrinsic variable of type RR Lyrae variable with a period of about 13.6 hours. First, spectra of RR Lyr were taken to cover one complete pulsation cycle of RR Lyr. Next, photometric data was taken of RR Lyr to cover one complete pulsation cycle and create a light curve. The spectra of RR Lyr were then taken into a spectroscopy software in order to determine their spectral type by comparing the depths and widths of the Hδ and Hγ spectral lines. Then, the photometric data was compiled into a light curve of RR Lyr. The hypothesis that as the brightness of RR Lyr increased, so would its temperature was supported by the findings of this experiment. It was found that there was a direct relationship between the brightness of RR Lyr and the temperature.

1. Introduction
The purpose of the project was to study the light curve of variable star RR Lyr, and determine if there was a correlation between the magnitude and the temperature. It was hypothesized that as the brightness of RR Lyr increased, so would its temperature. RR Lyr is a variable star of type RR Lyrae Variables. RR Lyr went under a spectroscopic and photometric analysis in order to test the previously mentioned hypothesis.

There are many types of variables falling into three major categories: eclipsing binaries, cataclysmic variables, and pulsating variables. Eclipsing binaries occur when two stars in orbit around each other periodically block the light of the other which results in periodic dips in the magnitude of the system. An eclipsing binary would not have been a good candidate for this project because the dips in brightness would have no correlation to a change in temperature. The second type of variable star is a cataclysmic variable. Cataclysmic Variables occur as a result of a violent event in a star. These variables are characterized by their sudden change in brightness. These variables would portray a temperature variation as well as a brightness variation, but are much harder to find, which makes them less preferable to the third type of variable star, the pulsating variable. A pulsating variable star is a star that varies in brightness because of the nature of the star itself, rather than as a result of being in a binary system. This type of variable is sometimes called an intrinsic variable because the star itself is causing the variance. A pulsating variable was the best candidate for the project because the time it takes to complete one full pulsation period has been well calculated and the variance can be studied over multiple nights. RR Lyrae variables have a period ranging from 0.5 to 1 days, making it the ideal selection for this project. The reason for the pulsation was first theorized by a British Astrophysicist named Sir Arthur Eddington in 1941. He suggested that the periodicity was due to the ease at which radiation can pass from the core of a star to its photosphere, its opacity. As the opacity of a star increases, the radiation becomes trapped and increases the pressure inside of the star causing it to “puff-up.” On the contrary, if the opacity falls, it would cause a decrease of internal pressure inside of the star because the radiation could escape more readily, decreasing the size of the star. According to Eddington, under certain circumstances, a star could become unstable and cause periodic risings and fallings of opacity in a star resulting in the pulsations that we see. These special conditions required to form a pulsating variable cannot be found in main-sequence stars, but only in a region on the H-R diagram known as the instability strip. The star used in this project was the first of its kind to be discovered, RR Lyr. RR Lyr was the first RR Lyrae variable to be discovered and has an ideally short period of 13.6 hours. RR Lyr exhibits an effect over a long period of time known as the Blazhko effect, a variance in period and amplitude over a long period of time. Along with RR Lyr, about 20% of all RR Lyrae variables exhibit the Balzho effect. After the
research, RR Lyr was determined to be the best star for the study.

Along with selecting a variable star for the study, it was necessary to find spectral lines that would be best suited for a comparison between RR Lyr and a known spectra in order to determine the spectral class, and from that, the temperature of RR Lyr at each point in the pulsation cycle. After examining a few spectral lines, and following Gray & Corbally (2009), it was determined that the Hγ and Hδ lines would be best suited for the study as they are not affected by any nearby spectral lines.

When doing differential photometry, many precautions need to be taken, one involves the focusing of the star. When focusing a star for photometry it is important to focus it so that the point spread function spreads over at least three pixels. This is done to minimize the amount of light lost in the gap between the individual pixels in the CCD which would cause a distortion in the magnitude difference during differential photometry. Stars are so far away from the observer that a perfect image of a star should appear as a single point because it has an infinitely small angular size. Unfortunately, when an image is taken of a star using a CCD camera, the star appears as a circle. This, almost blurring, of the light is due to many factors including diffraction of the electromagnetic wave at the finite aperture of the lens, design and fabrication tolerances in the surfaces of the lens, some optical aberrations that are virtually unavoidable in modest-cost consumer-grade optics, atmospheric turbulence, and an imperfect focus. All of these effects combined form the imperfect image in the CCD camera.

2. Data Collection

2.1 Spectroscopy

Setting up the spectroscopy instrumentation required the addition of a spectrometer, a CCD camera, and an autoguider to the telescope. Before an observing night began, a series of flat-field compensation images were taken using the tungsten lamp in the Alpy spectrograph. During each night’s data collection, spectra of HD 181470 were taken for wavelength calibration during data reduction. Spectra were also taken using the Ar-Ne lamp in the spectrometer in order to help calibrate for wavelength as well. More spectra were taken of HD 180163, a B2.5IV star, were taken for atmospheric and instrumental response calibration. After the calibration images were taken, series’ of RR Lyr spectra were taken as long as the conditions remained favorable. A total of four nights were devoted to spectroscopy, in order to cover the entire pulsation cycle.

2.2 Photometry

For the photometric data collection, a field including HD 182755 and RR Lyr was selected and many images were taken along one pulsation phase. The exposure times used for the B, V, and R filters were 20 seconds, 10 seconds, and 10 seconds respectively. A total of five nights were devoted to photometry, in order to map the complete lightcurve.

2.3 Calibration Images

In the spectroscopy setup, calibration images were taken for flat field, wavelength calibration, and atmospheric and instrumental response were taken as previously discussed. Dark and bias images were taken as well. One set of dark images and bias images were used for the data set. In the photometry setup, bias, dark, and flat frames were used from the observatory library, for reduction of the CCD images.

3. Data Reduction

3.1 Spectroscopic Data Reduction

The software ISIS was used for the spectroscopic data reduction. One of the calibration spectra of HD 181470 was used to calibrate for the smile and tilt of the spectral lines in the images. Calibration images of HD 180163 were used to calculate an instrument response curve that was calibrated between the wavelengths of 3800Å and 7000Å. The instrument response curve was then applied to each of the following spectra. The spectra of RR Lyr were opened after being calibrated with the dark, bias, flat-field, and instrument response curve. These spectra were then cropped to 4000Å and 5000Å because the Hγ and Hδ lines lie between these wavelengths. The spectral class of RR Lyr was determined by comparing the depths of the Hγ and Hδ lines to those of stars with a known spectral class.

3.2 Photometric Data Analysis

The software used for the differential photometry was MPO Canopus. The dark, bias, and flat frames were applied to the photometry images using CCD Soft. The reduced images were then opened in MPO Canopus and the differences in magnitude of RR Lyr and HD 182755 were calculated for each filter.
3.3 Composition of Photometric Data

When composing the photometric data, the current fraction of the pulsation cycle was calculated using the equation \( \phi = \frac{[\tau_1 - \tau_0]}{\rho} \). Each of the values for the magnitude of RR Lyr and its corresponding pulsation phase fraction were graphed yielding a light curve of RR Lyr to compare with the spectroscopic data.

4. Results

4.1 Light Curve of RR Lyr

4.2 Temperature Curve of RR Lyr

5. Analysis of Final Data

The results of the experiment supported the hypothesis that as the magnitude of RR Lyr increased, so would its temperature. The data showed that the temperature and magnitude of the star were almost directly proportional. As the magnitude increased, so did the temperature of RR Lyr. As shown in both the light and temperature curve, if a line is interpolated along the graph of temperature vs. pulsation phase, it almost exactly resembles the light curve of RR Lyr. The light curve of RR Lyr resembled the light curve that was expected, except for the small delay along the decrease in magnitude around pulsation phase 0.28. The results not only supported my hypothesis, but they also support Sir Arthur Eddington’s theory of why intrinsic variable stars pulsate. Eddington proposed that the pulsating was controlled by a rise and fall in the opacity of the star. His theory suggested that as the opacity rises, the radiation will become trapped and the star will expand, which causes a decrease in both luminosity and temperature. Also, he suggested that as the opacity decreases, the star will allow radiation to more readily escape and the star will shrink causing an increase in both luminosity and temperature. The data supports this because there is a direct relationship between the temperature of the star and its magnitude, just as Eddington’s theory proposed. Assuming that the opacity theory is correct, it would follow that the reason for the delay in the light curve of RR Lyr would be due to some mechanism hindering the decrease in opacity of RR Lyr. It was inferred that while the opacity was decreasing, something slowed down that decrease much like a catalyst can slow down a chemical reaction.

6. Conclusion

The purpose of this experiment was to compare the temperature of RR Lyr and its magnitude to determine if there was a relationship. In the experiment, two different studies were done on the variable star RR Lyr. The two studies were a spectroscopic analysis and a photometric analysis. A series of spectra were taken of RR Lyr to cover one entire pulsation phase. The hypothesis that as the magnitude of the star increased, so would its temperature was supported. It was also found that there was a direct relationship between the temperature of RR Lyr and the magnitude of RR Lyr, which supported a theory proposed by Sir Arthur Eddington in 1941 that the reason that stars pulsate is...
due to an increase and decrease in the opacity of the
star.

7. Acknowledgements

I am extraordinarily grateful towards Robert
Buchheim who mentored me through this project
and has provided the much needed help in
overcoming the learning curve that was required of
me during the project. Not only did he help with the
understanding of my data, but he allowed me to use
his personal observatory for the collection of the data.

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A Slitless Spectrograph That Provides Reference Marks

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Abstract
The author designed and built a slitless spectrograph to record reference marks along the spectrum of a point light source. Spectra can be taken of transient, clustered, or moving lights when a spectrograph cannot be accurately aimed at the lights to capture slit spectra. Three beams of undispersed light, directed by mirrors and lenses, provide reference marks. Near each end of the spectrum a reference mark barely varies from the corresponding point on the spectrum when the aim toward the light source varies. Within 2 degrees of perfect aim toward the light source, the variation is less than 7 angstroms. The third reference mark enables this variation to be quantified. The locations and orientations of the optical components are mathematically derived. Additional features of the spectrograph enable the use of a slit and comparison spectrum, and the recording of higher orders by moving the camera and using specific Wratten filters.

1. Introduction
The need for this spectrograph became evident to the author during his pursuit of a hobby studying nocturnal luminous phenomena. The mysterious Brown Mountain light of North Carolina was the primary target. Appalachian State has been investigating these lights (Appalachian State). The following versatile characteristics were desired in the design of the spectrograph:
1. Reference marks in the slitless mode for short-lived, moving, and clustered lights, when precise aim toward target is not possible for use with a slit spectrograph and a comparison spectrum.
2. Ability to record entire visible spectrum at one time in first order.
3. Fast optics.
4. Highest feasible resolution.
5. Portability by hand to a site some distance from the nearest parking area.
6. Self-contained, not attached to a telescope.
7. Comparison spectrum while operating in the slitless mode.
8. Portability in a suitcase aboard an airplane.
9. Ability to record higher orders by means of filters and various positions of the camera.
10. Automatic frame advancer for recording fast-changing phenomena.

2. Previous Methods of Providing Reference Marks for Slitless Spectra
In a commonly-used method, the undispersed or "zero order" light from the source is used for a reference mark, although recorded far from the spectrum, as reported by Liu (1962) and Remus (1978).

Various methods have been proposed for providing spectral reference marks on the spectrogram recorded by a slitless spectrograph. One category of methods consists of inducing dark lines on the spectrum. One method uses transparent plane-parallel glass plates within the optical path for recording interference bands as reported by Linnik (1971). Another method records absorption lines generated by an absorbing medium in the optical path as reported by Fehrenbach (1947, 1966) and Millman (1931). Griffin (1973) reported still another method using telluric absorption lines.

Another category of methods uses convex cylindrical surfaces both to reflect the source light as a narrow image and to reflect a laboratory source of light for comparison as reported by Baskett (1970) and Lebedinsky (1955). These methods are used
primarily for taking spectra of extended illuminated areas.

Still another category of methods uses a second beam from the light source to provide either a relatively undispersed point of light along the spectrum or a second spectrum. In one method, the spectrum of a star is simultaneously photographed next to a reversed spectrum of the same star as described by Fehrenbach (1966, 1947a), Millman (1931), Geyer (1979), Gieseking (1979), and Comstock (1906).

Kirillovykh (1980) reports another method using a prism and light filters in front of a diffraction grating to produce reference points on the spectrum.

3. Description of the Spectrograph

The method described in this paper uses redirected light beams from the source to provide accurate and correctable reference marks as previously described by Buchanan (1984).

Figure 1 is a diagram of the subject slitless spectrograph.

Figure 2 is a diagrammatic plan view of the beams of light passing through the diffraction grating and camera lens to the film.

Figure 3. Enlarged View of Small Portion of Grating Surface, Showing Detail of Incident and Emerging Beams
Figure 3 shows an enlarged view of a small portion of the grating surface, showing detail of the incident and emerging beams.

Ideally, the light source is a point sufficiently distant so that all photons from it are essentially parallel. A spotting telescope containing cross hairs is used to align the spectrograph toward the source. The optical elements consist of plane mirrors, lenses and a diffraction grating. All optical elements are mounted in the plane which is perpendicular to the grooves of the diffraction grating. These elements are positioned to direct four beams of light through a camera lens which focuses the light onto a recording medium. Although various recording media, such as photosensitive film or a charge couple device can be used, film is assumed in this paper.

The first beam of light strikes the diffraction grating, which disperses the light, then passes through a lens which focuses the spectrum on the film. The first beam makes an incident angle “i” with the normal to the surface of the diffraction grating, is dispersed by the grating, and emerges at an angle “S” with respect to the axis of the camera lens.

The second beam strikes a plane mirror which reflects the beam onto another plane mirror which reflects the beam through the grating, from which the camera lens. The undispersed or zero order light passes through the lens which focuses the light onto the film at a selected position on the spectrum, forming the first reference mark.

The third beam strikes a plane mirror which reflects the light onto another plane mirror which reflects the light through a negative lens and a positive lens, resulting in collimated light at a selected magnification. This pair of lenses can also be placed in front of the mirrors. The third beam emerges from the grating at an angle “B” from this axis. The undispersed light passes through the grating and the lens, which focuses the light onto the film at a selected position on the spectrum, forming the second reference mark.

The fourth beam is reflected by a mirror directly through the grating. The lens focuses the undispersed light of the fourth beam on the film to form a third reference mark at a varying position on the spectrum. The fourth beam emerges at an angle “W” from this axis.

4. Equations for the Four Light Beams and the Optical Constants

The formulas and numerical constants of the spectrograph are derived. The number of digits used is for mathematical demonstration, and does not imply physical precision. Angles are in radians unless noted otherwise. All terms calculated with subscript “O” are constants, referring to the condition when the spectrograph is perfectly aimed toward a light source.

4.1 First Beam

The equation of diffraction of light by the grating for a first order spectrum is

$$\frac{\lambda}{g} = \sin i + \sin (S + Y)$$

where “λ” is the wave length of a spectral line in a dispersed light beam. The term “Y” is the angle between the normal to the surface of the diffraction grating and the axis of the camera lens. The term “g” is the groove spacing on the diffraction grating. A grating of 600 grooves per mm is selected. The groove spacing is

$$g = \frac{10^7 \lambda}{600} = 16,666.667 \lambda$$

4.2 Second Beam

The first and second reference marks are positioned near the opposite ends of the first order of the visible spectrum. The first mark “λ1” is set at the B-solar line due to oxygen, according to The Chemical Rubber Company (1967),

$$\lambda_1 = 6,869.955 \text{ Å}$$

When the spectrograph is perfectly aimed,

$$R_0 = S_0$$

where “S_0” is the angle of the diffracted first beam with respect to the axis of camera lens at the first reference mark.

To make angle “R” of the undiffracted second beam to change at the same absolute rate as angle “S”, when the angle “i” changes,

$$\frac{dR}{di} = \frac{dS}{di} = -1$$
To differentiate equation (1) with respect to “i”, when “λ”, “g”, and “Y” are constant,

\[ 0 = (\cos i) \frac{di}{dS} + \cos (S + Y) \cdot dS \]  

(6)

\[ \frac{dS}{di} = - \frac{\cos i}{\cos (S + Y)} \]  

(7)

When “i₀” and “S₀” are substituted,

\[ \cos (S₀ + Y) = \cos i₀ \]  

(8)

Solution 1: \( i₀ = S₀ + Y \)  

(9)

Solution 2: \( i₀ = -(S₀ + Y) \)  

(10)

Solution 1 applies for the diffracted first order of the first beam, the condition desired. Solution 2 applies to the zeroth order of the first beam, and is of no further interest.

To solve for angle “i₀”, the incident angle of the first beam at precise aim, appropriate constants are substituted into equation (1),

\[ \frac{\lambda₁}{g} = \sin i₀ + \sin (S₀ + Y) \]  

(11)

Since \((S₀ + Y) = i₀\), from equation (9),

\[ \frac{\lambda₁}{g} = 2 \sin i₀ \]  

(12)

\[ \sin i₀ = \lambda₁ / 2g = 0.20609865 \]  

(13)

\[ i₀ = 0.20758633 \text{ radian} = 11.893821° \]  

(14)

### 4.3 Third Beam

The second reference mark “\( \lambda₂ \)” is set at a wavelength of HI, according to the Massachusetts Institute of Technology (1969),

\[ \lambda₂ = 3889.055 \text{ Å} \]  

(15)

This wavelength is also known as hydrogen-zeta. To make the two reference marks to be equally spaced in angle from the axis of the camera,

\[ S₀ = -S'₀ \]  

(16)

where “-S'₀” is the angle of diffracted first beam with respect to axis of the camera lens at the second reference mark. Substituting into equation (1),

\[ \lambda₁ \frac{1}{g} = \sin (i₀) + \sin (Y + S₀) \]  

(17)

and

\[ \lambda₂ \frac{1}{g} = \sin (i₀) + \sin (Y + S'₀) \]  

(18)

or

\[ Y + S'₀ = Y - S₀ = \arcsin [(\lambda₂ / g) - \sin (i₀)] = 0.027248022 \text{ radian} = 1.5611967° \]  

(19)

To subtract equation (17) from equation (9),

\[ Y + S₀ - (Y - S₀) = 2 \cdot S₀ = i₀ - 0.027248022 \]  

(20)

\[ S₀ = 0.09016915 \text{ radian} = 5.1663117° \]  

(21)

\[ S'₀ = -0.09016915 \text{ radian} \]  

(22)

\[ Y = i₀ - S₀ = 0.11741718 \text{ radian} = 6.727509° \]  

(23)

To solve for “R₀” by substituting in equation (1),

\[ \lambda₁ \frac{1}{g} = \sin i₀ + \sin (R₀ + Y) \]  

(24)

\[ R₀ = 0.09016915 \text{ radian} = 5.166312° \]  

(25)

To solve for “B₀” by substituting in equation (1),

\[ \lambda₂ \frac{1}{g} = \sin (i₀) + \sin (Y + B₀) \]  

(26)

\[ B₀ = -0.09016915 \text{ radian} = -5.166312° \]  

(27)

When the spectrograph is precisely aimed,

\[ S'₀ = B₀ \]  

(28)

The angle “B” of the undiffracted second beam is selected to change at the same rate as angle “S” of the diffracted first beam, when angle “i” changes. This condition can be satisfied only if the third beam is magnified.
\[ \frac{dE}{di} = 1 \]  
\[ \frac{dB}{di} = m_0 \left( \frac{dE}{di} \right) = \frac{dS}{di} \]  
\[ \frac{dS}{di} = -m_0 = -\cos(i_0) / \cos(S'_0 + Y) = [\cos(i_0)] / \cos(S'_0 + Y) \]  
\[ m_0 = 0.9788946 \]  

4.4 Fourth Beam

The fourth beam emerges from the grating at an angle “\(W\)”, which changes in the opposite direction from the emerging diffracted beam when the angle of incidence changes. The fourth beam is reflected once and is unmagnified. To determine “\(W_0\)”, we set the fourth beam to be midway between the second and third beams when “\(i = i_0\)”. Therefore,

\[ W_0 = \frac{(R_0 + B_0)}{2} = 0 \]  
\[ \frac{dW}{di} = -\frac{dR}{di} = \frac{1}{m_0} \left( \frac{dB}{di} \right) = 1 \]  

Therefore,

\[ R - R_0 = W_0 - W = \frac{(B - B_0)}{m_0} = i_0 - i \]  

5. Positions of Reference Marks on Film

Figure 4 is a diagrammatic plan view showing the camera lens, the emerging angles of the undiffracted second, third, and fourth beams, the film, and the corresponding reference marks when the spectrograph is precisely aimed. The dimension “\(f\)” represents the focal length of the camera lens, which is selected to be 135 mm.

\[ f = 135 \text{ mm} \]
Figure 5. Diagrammatic Plan View of Camera Lens, Emerging Undiffracted Beams, Film, and Reference Marks When Spectrograph is Not Precisely Aimed at Unknown Light Source.

From Figures 4 and 7 these equations can be deduced:

\[ a = f (\tan R - \tan W) \]  
\[ b = f (\tan W - \tan B) \]

and that specific equations for perfect aim are:

\[ a_o = f (\tan R_o - \tan W_o) = 12.205934 \text{ mm} \]  
\[ b_o = f (\tan W_o - \tan B_o) = 12.205934 \text{ mm} \]

6. Derivation of Formula for Calculating an Unknown Spectral Line

Figure 8 is a diagrammatic representation of a spectrogram recorded when the spectrograph is aimed generally toward a group of three light sources having various elevations and directions. The dimension “d” is the distance from the first reference mark to a spectral line. When the spectrograph records multiple point sources of light, the spectra of all sources along with their respective reference marks will be simultaneously recorded.

Subtracting Equation (21) from Equation (19),

\[ (a - a_o) = f (\tan R + \tan W_o - \tan R_o - \tan W) \]

When substitutions for “R” and “W” are made from equation (17),

\[ (a - a_o) / f + \tan R_o - \tan W_o = - \tan [(i - i_o) - R_o] - \tan [(i - i_o) + W_o] \]  
\[ (a - a_o) / f + \tan R_o = - \tan [(i - i_o) - R_o] \]

Since \( W_o = 0 \), from equation (20),

\[ a_o / f = \tan R_o \]

and

\[ a / f = -\tan [(i - i_o) - R_o] \cdot \tan (i - i_o) \]

where “a” is the independent variable and “i” is unknown. Separating the “i” term and placing all terms containing “i” on the left side of the equation:

\[ \tan [i - (i_o + R_o)] + \tan (i - i_o) = - a / f \]

To use this trigonometric identity,

\[ \tan (x - y) = (\tan x - \tan y) / [1 + (\tan x)(\tan y)] \]

\[ \frac{\tan i - \tan (R_o + i_o)}{1 + (\tan i)(\tan R_o + i_o)} \]

\[ \frac{\tan i - \tan i_o}{1 + (\tan i)(\tan i_o)} = \frac{- a}{f} \]  
\[ \frac{(\tan i - \tan i_o)}{1 + (\tan i)(\tan i_o)} = \frac{a}{f} \]  
\[ \frac{(\tan i - \tan i_o)}{1 + (\tan i)(\tan i_o)} = \frac{a}{f} \]

To facilitate algebraic manipulation, the following terms are substituted:

\[ l = \tan i \]

where “l” is a dependent variable to be solved.

\[ J = \tan (i_o) = 0.21062041 \]
\begin{align*}
L &= \tan (R_0 + i_0) = 0.30687866 \\
\text{Therefore,} \quad (I - L) / (1 + IL) + (I-J) / (1 + IJ) &= - a / f \\
\text{Using the common denominator,} \\
[ (I - L) (1 + IL) + (I - J) (1 + IL) ] / [ (1 + IL) (1 + IJ) ] &= - a / f \\
\text{Expanding terms and multiplying equation through by: "} f (1 + IL)(1 + IJ) \text{"}, \\
\text{If} + I^2 J f + J f - L f - I^2 J f - J f - I J f &= -a \left[ J + L + (I^2) J L \right] \\
\text{Collecting terms in descending exponents of "} I \text{"}, \\
I^2 [f(J + L) + (JL)a] + I [2f(1 - JL) + (J + L) a] + [-f(J + L) + a] &= 0 \\
A_q I^2 + B_q I + C_q &= 0 \\
\text{Since the solution to the quadratic equation is} \\
x &= \frac{-b \pm \sqrt{b^2 - 4ac}}{2a} \\
I &= - \frac{B_q \pm \sqrt{[B_q]^2 - 4 A_q C_q^{0.5}}}{2A_q} \\
\text{Setting} \ A_q = 1/2, \\
I &= - \frac{B_q \pm \sqrt{[B_q]^2 - 2C_q^{0.5}}}{2A_q} \quad (26) \\
\text{To set} \ A_q = 1/2, \text{ divide the a-equation through by "} 2[f (J + L) + JLa] \text{"}, \\
(I^2) / 2 + \frac{-f(J + L) + a}{[2f (J + L) + 2JLa]} &= 0 \quad (27) \\
\text{Put "} a \text{"} \text{ terms on left and divide numerator and denominator of "} B_q \text{" by "} 2JL \text{"}, \\
B_q &= \frac{[a (J + L) / (2JL) + f (1 - JL) / JL]}{a + f (J + L) / JL} \\
\text{We divide denominator into numerator, resulting in only one input term "} a \text{"}, \\
B_q &= \frac{J + L \cdot f/JL / [J + L]^2 / 2[2JL] / \sqrt{[J + L]}}{a + f(J + L) / JL} \quad (28) \\
\text{Quantifying all known terms,} \\
B_q &= 4.0032474 \\
- \left[ 2373.3627 / (a + 1080.8768) \right] \quad (29) \\
\text{Put "} a \text{"} \text{ terms on left and divide numerator and denominator of "} C_q \text{" by "} 2JL \text{"}, \\
C_q &= \frac{(a / 2JL) - (f / 2JL) (J + L)}{a + f (J + L) / JL} \quad (30) \\
\text{We divide denominator into numerator, resulting in only one input term "} a \text{"}, \\
1 / (f / 2) (J + L) &\left[ (1 / JL^2) + (1 / JL) \right] \\
C_q &= \frac{2JL - a + f (J + L) / JL}{2JL} \quad \text{Quantifying all known terms:} \\
C_q &= 7.735758 - \left[ 8901.839 / (a + 1080.8768) \right] \quad (31) \\
\text{To compare actual measurement of "} (a + b) \text{"} \text{ against theoretical value, we solve for "} (a + b) \text{"} \text{ as a function of "} i \text{"} \text{ from Equations (16) and (19),} \\
a + b &= a + f \{ \tan [(i - i_0) + W_0] - \tan [B_0 - m_0 (i - i_0)]) \} \quad (32) \\
\text{Since} \ W_0 &= 0, \\
a + b &\text{ } \text{ } \text{ } \text{ } \text{ } \text{ } \text{ since} \ W_0 &= 0, \\
\text{a + b} &\text{ } \text{ } \text{ } \text{ } \text{ } \text{ } \text{ a + f} \{\tan [(i - i_0)] - \tan [B_0 - m_0 (i - i_0)])\} \\
a + 135 &\tan [I - 0.20758633] + \tan [0.9788946 (I - 0.20758633)] + 0.09016915] \quad (33) \\
\text{7. Calculation of an Unknown Spectral Line} \\
The spotting scope is aimed imprecisely at an unknown light source, and an exposure is made. Measurements on the developed film disclose that: \\
a &= 19.237 \text{ mm}, \\
b &= 5.265 \text{ mm}, \text{ and} \\
d &= 4.856 \text{ mm} \\
\text{From Equations (25, 26, 29, 30),}
i = 0.18168767

From Equation (17),

\[ R = i_0 - i + R_0 = 0.11606781 \]

\[ h = f \tan R \] (34)

where “h” is the distance of the first reference mark from axis of camera lens.

\[ h = 135 \tan R = 15.739899 \]

\[ j = -f \tan B \] (35)

where “j” is the distance of second reference mark from axis of camera lens.

The term “Su” is the angle of the unknown spectral line with respect to axis of camera lens:

\[ \tan Su = \frac{h - d}{f} \] (36)

\[ Su = \arctan \left[ \frac{(h - d)}{135} \right] = 0.08044748 \]

The wavelength of the spectral line, “\( \lambda_u \)”, is determined by substitutions in equation (1),

\[ \lambda_u = g \left[ \sin i + \sin (Su + Y) \right] = 6287.8 \text{ Å} \] (37)

8. Spreadsheet Solution

Since this slitless spectrograph was first reported (#15), spreadsheet software for personal computers has become widely available. The term "i" is solved for each value of "a". See TABLE 1.

The formulas in each column are as follows:

<table>
<thead>
<tr>
<th>COLUMN</th>
<th>ITEM</th>
<th>FORMULA</th>
</tr>
</thead>
<tbody>
<tr>
<td>1</td>
<td>d</td>
<td>input data</td>
</tr>
<tr>
<td>2</td>
<td>a</td>
<td>input data</td>
</tr>
<tr>
<td>3</td>
<td>Bq</td>
<td>Equation (29) w/ COL2 input</td>
</tr>
<tr>
<td>4</td>
<td>Bq squared</td>
<td>(COL3)exp(2)</td>
</tr>
<tr>
<td>5</td>
<td>Cq</td>
<td>Equation (31) w/ COL2 input</td>
</tr>
<tr>
<td>6</td>
<td>I</td>
<td>-COL3+[COL4-2*(COL5)]exp(.5)</td>
</tr>
<tr>
<td>7</td>
<td>i</td>
<td>arctan(COL6)</td>
</tr>
<tr>
<td>8</td>
<td>tan W</td>
<td>tan(COL7-i0)</td>
</tr>
<tr>
<td>9</td>
<td>tan B</td>
<td>tan[m0*(COL7-i0)]+B0</td>
</tr>
<tr>
<td>10</td>
<td>a + b</td>
<td>COL2+[135*(COL8+COL9)]</td>
</tr>
<tr>
<td>11</td>
<td>h</td>
<td>135*tan(i0)-COL7+B0</td>
</tr>
<tr>
<td>12</td>
<td>su</td>
<td>arctan[(COL11-COL1)/135]</td>
</tr>
<tr>
<td>13</td>
<td>( \lambda )</td>
<td>g*[sin(COL7)+sin(COL12+Y)]</td>
</tr>
<tr>
<td>14</td>
<td>a/(a+b)</td>
<td>COL2/COL10</td>
</tr>
<tr>
<td>15</td>
<td>error</td>
<td>COL13-(either ( \lambda_1 ) or ( \lambda_2 ))</td>
</tr>
</tbody>
</table>

The term "d" is first input at COL12; therefore, for earlier columns, the values on lines 2-12 are repeated from line 13 downward. They are omitted in COL3-COL6 and COL8-COL10 since they are not used in subsequent calculations.

TABLE 1 shows that the error between the first reference mark and the target wavelength position ranges up to 7.0 Å. The error between the second reference mark and the target wavelength position ranges up to 3.9 Å. FIGURE 9 shows the errors as related to a/a+b.

From TABLE 1, the average dispersion for i = 0.2065, (which roughly approximates i0) is:

\[ \frac{6869.955 - 3889.055}{24.415} = 122 \text{ Å/mm} \]

Since the readings are expressed to the nearest 0.001 mm, the corresponding precision of wave length is:

\[ 122 \text{ Å/mm} * 0.001 \text{ mm} = 0.122 \text{ Ångstrom}, \text{ rounded off to 0.1 Ångstrom}. \]

The width of exposed film for 35 mm film is about 36 mm. The half length of exposed area on film is:

\[ 36 \text{ mm}/2 = 18 \text{ mm}. \]

When "a" = 0, first and third reference marks coincide. From Eq. (35),

\[ j = -f^*\tan B = -f^*\tan[-m0(i - i0) + B0] = 18.240 \text{ mm} > 18 \text{ mm}; \]

the second reference mark falls outside the exposed area.

When "a" = 24.5 mm, and "a + b" = 24.591 mm, second and third reference marks almost coincide.

From TABLE 1, h = 18.39 > 18 mm; the first reference mark falls outside the exposed area.
Therefore, when all three reference marks are present, the third mark always lies between the first and second marks.

9. Description of Components in the Constructed Prototype

The wooden housing of the spectrograph measures 66 cm by 33 cm by 17 cm and weighs 7.3 kg, including 35-mm camera and lens. The spectrograph can be operated in either the slitless mode or the slit mode. A separate unit weighing 2.5 kg and containing an objective lens, a slit, and a collimating lens can be inserted into the spectrograph housing, and a comparison spectrum can be exposed adjacent to the target spectrum.

The dispersing element of the prototype is a blazed transmission diffraction grating having 600 grooves per millimeter with a ruled length of 76 mm and a ruled width of 65 mm. The lens and film are parts of a commercial 35 mm camera. The camera lens has a focal length of 135 mm and an aperture of 71 mm. An automatic film winder is attached to the camera.

All mirrors are plane first surface mirrors to avoid ghost reflections. To eliminate interference with the spectrum, one of the mirrors in each optical train may be tilted minutely to reflect the beam to a point displaced a small vertical distance from the spectrum.

Spherical lenses in the third beam will cause a slight shift of the vertical position of the second reference mark respect to the spectrum of a light displaced from the horizontal axis of the system. Cylindrical lenses magnify light only in the direction perpendicular to the grating grooves. Therefore, cylindrical lenses are preferred in order to limit the distance of shift. Also, when the incident light varies vertically from the horizontal axis of the system, the reference marks will vary somewhat in vertical position with respect to the spectrum due to the distortion of the spectrum, according to Miranda (1965).

10. Additional Features of the Spectrograph

The spectrograph contains additional features for versatility. A hinged flap can be lowered to cover the paths of the second, third, and fourth beams to prevent reference marks from being recorded. The reference marks may be recorded during some portion of a long exposure. A removable unit that contains an objective lens, a slit, and a collimating lens can be inserted for operation in the slit mode. These lenses are identical to the camera lens. The unknown light focused by the objective lens illuminates one half of the slit; a known emission source shines onto a mirror which reflects the light through the other half of the slit. The known emission source for the slit is a small fluorescent light which produces spectra of mercury and argon.
Figure 9. Front Top View of Spectrograph

Upper left area contains optical elements to provide reference marks along slitless spectra. Hole in bottom allows light from calibration lamp to strike mirror on slit to form a comparison spectrum.

Figure 10. Top Rear View of Spectrograph

Camera in Position 3
The camera lens with camera and film are mounted on a board which can be locked at any of several positions about an axis vertically aligned with the center of the diffraction grating. The first order of the visible spectrum can be photographed in one position, and parts of the second and third orders can be photographed in other positions. Various Wratten filters are used to prevent overlapping of orders. The filters are described by Eastman Kodak Company (1973).

Spectrograph assembled before lid is placed. Unit containing objective lens, slit, and collimating lens is in place. The corner of camera support is set on a dot, position #1, for first order. Seven dots mark positions for recording sections of second and third orders.
Figure 13. Inside Spectrograph Showing Slit Unit

Figure 14. Orion Belt

Figure 15. AlpOri
11. Conclusions

This slitless spectrograph has particular applicability to determining spectra of moving, multiple, and transient luminous events. The film version requires no outside power source or internal light. The operator roughly points the instrument at a light source and takes a photograph. The practical field width approaches 2 degrees in every direction from the light source.

12. Acknowledgments

W. R. Buchanan, my father, in assisting the construction of the spectrograph.

The drafter, whose name is lost, who drew the original Figures 2-8 in India ink on 100% cotton paper in 1985.

J. R. Washburn, who primed my interest in spectroscopy while we were college students.
13. References

Appalachian State, (http://brownmountainlights.org/)


\[
\lambda / g = \sin i + \sin (S + Y),
\]

\[
\lambda = \sin i + \sin (S + Y) \quad (1)
\]

\[
g \quad = \quad 10^7 \lambda 
\]

\[
g = \frac{10^7 \lambda}{600} \quad = \quad 16,666.667 \lambda \quad (2)
\]