A SEARCH FOR LONG-PERIOD VARIABLE STARS IN THE GLOBULAR CLUSTER

NGC 6496

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ABSTRACT

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Studying the late stages of stars is essential to understand the process of stellar evolution. Depending on their masses and properties, some stars become unstable at the end of their evolutionary state and hence they start pulsating. Their brightness and surface temperature change with their pulsations and hence we see them as variable stars. We are looking for long period variable stars (LPVs) in the globular cluster NGC 6496.

We observed the cluster from February 2009 till October 2010 using a 0.41m telescope in the V and I bandpasses. We have identified 11 variable stars in the cluster. 6 of them are new discovered LPVs in which 3 of them are semiregular LPVs and the rest 3 are irregular LPVs. We plotted the color magnitude diagram (CMD) of this cluster and all our LPVs were detected on the RGB/AGB. 5 of the 11 variable stars are short period variable stars in which 4 of them are W UMa binary stars and 1 is an Algol binary star. The light curves of all these stars are plotted in this paper and the periods were detected using different period-finding methods.
My work is dedicated to my Dad and Mom

who offered me unconditional love and support throughout my life.
ACKNOWLEDGMENTS

This research project would not have been possible without the support and help of my advisor, Dr. Andrew C. Layden. Dr. Layden was abundantly helpful and offered all means of help and support and he was always available to direct and guide me. Deepest thanks are also due to the members of the supervisory committee, Dr. John B. Laird and Dr. Dale W. Smith without whose knowledge and assistance this thesis would not have been successful.

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I am heartily thankful to my siblings, Tania, Lara, Katia, Hussein and Aya whose encouragement was essential to fulfill my dreams. Special thanks goes out to my brother, Dr. Hussein Abbas who was the first to encourage and help me perusing my master’s degree in the US. “Like branches on a tree we grow in different directions yet our roots remain as one.”

I thank one more individual I was fortunate to have met during my study time at Bowling Green State University, “Katlyn Hade” who showed me love and care and who was always there for me when I needed support. Katlyn has been my pillar, my joy and my guiding light, and I thank her for her love.

“Knowledge exists potentially in the human soul like the seed in the soil; by learning the potential becomes actual.” Imam Al-Ghazali
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CHAPTER 1
INTRODUCTION

Studying globular clusters (GCs) is essential for understanding the process of stellar evolution. GCs are made up of collection of stars that relatively have the same ages and metallicities. That is because stars in the same GC are usually formed from the same cloud of gas and at the same time. Studying stars with relatively the same metallicities, ages and distances is of great advantage for astronomers.

Metallicity is a measure of the proportion of “heavy elements” or “metals” that a star contains. It is usually expressed as [Fe/H], the log of the ratio of iron to hydrogen in the star divided by the iron to hydrogen ratio in the sun. In addition to that, usually GCs are free of gas and dust. This means that the light coming from most of the stars in a GC will face the same amount of reddening (although this is not always the case), another advantage in studying GCs. Because GCs are usually old, they consist of population II stars, which have lower metallicities when compared to population I stars. Population II stars reach the end of their evolutionary state and lose mass as gas and dust. Hence, this gas and dust will enrich the environment and will enrich the metallicity of newborn stars (population I). Hence, the metallicity of an astronomical object may provide an indication of its age of formation.

Since the metallicity of a new star is dependent on the death of older stars, it is important to study the late phases of evolution of stars. The fate of every star mainly depends on its mass and metallicity so the late phases of evolution of stars can be different from one star to another. When the star is in its main sequence stage, it undergoes fusion of hydrogen into helium within its core. The rate at which it does this and the amount of fuel available depends upon the mass of the star. When the star runs out of hydrogen in its core and an inert helium core is building up, it climbs up the red giant branch (RGB). After that, the star reaches the horizontal branch (HB) where it is powered by helium fusion in the core (via the triple-alpha reaction) and by hydrogen fusion in a shell surrounding the core. When the
helium core gets exhausted, the star ascends the asymptotic giant branch (AGB). Depending on their masses, some stars end up forming a planetary nebula with the core becoming a white dwarf. Other, massive stars end up with a supernova explosion.

Since our research focuses on one type of variable stars, long period variable stars (LPVs), we need to understand the difference between different kinds of variable stars in order to be able to classify them correctly.

1.1 Classifying a variable star

Variable stars can be divided into two categories: intrinsic and extrinsic. The variability of intrinsic variables comes from the changes of the internal physical properties of the star itself (temperature, radius, brightness, etc.). Intrinsic variable stars can be classified into three categories as well: pulsating variables, eruptive variables, and explosive variables. This study focuses on the pulsating variables. By contrast, the variability of extrinsic variables comes from external physical properties like rotation and eclipses. The two kinds of extrinsic variables are called eclipsing variables and rotating variables.

1.1.1 Extrinsic Variables

Eclipsing Variables

Eclipsing variables are stars that are paired by gravity and are in close orbit around each other. When the orbital plane of these stars is inclined towards the earth then we can see the light level dip as one star eclipses (passes in front of) the other [34].

One type of eclipsing variables is the W UMa stars. W UMa stars are contact binary stars that are close to each other and they usually share their gaseous envelopes. These two stars are constantly eclipsing each other because they are gravitationally distorted by one another. We detect the change in brightness of these stars when the stars are eclipsing each other. The two stars are usually equal in size and luminosity, with surfaces relatively close
in distance that occult each other during the course of revolution [8]. The components are of
different mass, but share very similar temperatures. Variability of these stars ranges from a
few tenths to slightly over a magnitude. The periods are typically short and range between
0.25 days to around 1.0 days.

Rotating Variables

Rotating variable stars are variable stars whose variability is caused by axial rotation
with respect to the observer. The nonuniformity of surface brightness may be caused by the
presence of spots or by some thermal or chemical inhomogeneity of the atmosphere caused
by a magnetic field whose axis is not coincident with the rotation axis [7].

1.1.2 Pulsating Stars (Intrinsic Variables)

There are two main types of pulsating stars: LPVs and short period variables:

Long Period Variable Stars

LPVs are pulsating stars that are not in hydrostatic equilibrium [30]. This causes the
star’s radius to swell and shrink regularly, sometimes with a well defined period (Mira),
sometimes semiregularly (SR) with an average period and amplitude and sometimes with an
irregular period (IRR). This pulsation instability causes the star to lose some of its mass as
gas and dust through stellar winds (that will later be used to form another star) [29]. (1) Mira
variables have long, rather regular pulsation cycles, with periods longer than 30 days and
large light variations, from 2.5 up to 11.0 magnitudes. (2) Semiregular variables have periods
in the same range as Miras but with lower amplitudes and semiregular periods which make
the period determination hard. (3) Irregular variables, similar to the semiregulars, are hard
to determine the periods for because of the irregularity of the light change. The irregularity
is more extreme than in SRs, both in terms of variations in period and amplitude from one
cycle to the next.
Based on the mentioned properties of variable stars (color, period’s length, period’s regularity, amount of the change of the magnitude and star’s position on the CMD), we can classify the type of the LPVs.

**Short-Period Variable Stars (Cepheids and Cepheid-like stars)**

The other type of pulsating stars are the short-period variable stars. They have short periods (between 1 and 30 days for Cepheids and 0.2 to 1 day for RR Lyrae variables) and their luminosity cycle is very regular. The majority of the RR Lyrae stars are found in the spherical component of the Galaxy and they pulsate while they are on the horizontal branch [7]. There are different types of RR Lyrae stars. The most common types are (1) RRab stars that have large amplitude, and an asymmetric light curve and their periods tend to be toward the long end of the RR Lyrae range. RRab stars pulsate in the fundamental mode. (2) RRc stars have variations in brightness due to their pulsations which are more sinusoidal and smaller in amplitude, and their periods tend to be toward the short end of the RR Lyrae range [34]. RRc stars pulsate in the first harmonic mode.

Finally, Cepheid variables are larger, more massive and in most cases younger than RR Lyrae stars and they have a longer periods. There are two main kinds of Cepheid variables: (1) Classical Cepheids and (2) Type II Cepheids. (1) Classical Cepheids undergo pulsation with very regular periods and they are population I variables stars. They are supergiant stars with masses more than 4 times the mass of the sun. (2) Type II Cepheids are population II stars (older than classical Cepheids), which are metal-poor and low mass objects. Although our research focuses on LPVs, we were able to detect 5 possible short period variable stars. The pulsations of the LPVs are detected by a variation of brightness and surface temperature over a long period of time. Understanding the properties of these pulsating stars is important for the process of stellar evolution and the process of enrichment of the universe because they are at a late phase of evolution.
Because of the irregularity of IRRs and SRs and because of their long periods, we have difficulty finding their periods. Although these stars do have specific periods, they have trouble expressing their periods consistently. In our study, our ability to find the period is limited because sometimes we are unable to cover many cycles for the star due to their long periods. Observing these stars for more cycles will help us getting more accurate periods and will also help us classifying these stars more precisely.

It is believed that as the LPVs climb on the RGB, they move from being IRR to SR and finally they become Miras. In other words, they pulsate in their fundamental mode when they are Mira stars, in third and second overtone when they are SRs and IRRs, respectively [3]. This classification is not confirmed, that is another reason to conduct research like ours. We can test this idea by plotting the period-luminosity (PL) relationship of our LPVs comparing it with the theoretical models. Also, understanding the properties of these pulsating stars will help astronomers improve the models of stellar evolution and pulsation and they will help us find the relationship between mass loss and stellar evolution [6].

In addition, theoretical isochrones can be compared to our CMD to see if they match our observations based on the specific age, distance and metallicity (see Chapter 5). If the models match our observations then we can confirm that the models are realistic and thus these models can be used for other projects. Finally, understanding LPVs in GCs will also help us understanding the properties of the last evolutionary stages of population II stars.

Most of the previous studies about variable stars focused on short period variable stars due to the periodicity of their light curves and due to their short periods, which are easier for astronomers to observe. For instance, to get the period of a RR Lyrae star it is sufficient to observe it for few hours. However, it takes from forty to thousands of days of observations if we want to get an accurate period of an LPV. This is a major reason LPVs haven’t been studied widely.

Moreover, most of the previous studies of variable stars in GCs focused on low-metallicity
GCs. Metal rich GCs have received less attention, in part because as members of the bulge or thick disk, most of these clusters lie at low galactic latitude and suffer from high and differential reddening and high contamination by field stars [1].

1.1.3 NGC 6496

NGC 6496’s properties have been carefully studied. Richtler et al. [22] observed NGC 6496 in August 1990 on La Silla with B and V filters. Another study of NGC 6496 was conducted by Armandroff [21]. Armandroff observed the cluster in V, R and I filters. Richtler et al. [22] and Armandroff [21] provided magnitudes on a standard system, which we can use to calibrate our photometry. Unlike Richtler et al., Armandroff plotted the actual color-magnitude diagram (CMD) of the cluster and estimated the amount of reddening of the cluster. Harris [24] and Stetson [19] also studied this cluster and determined many of the cluster’s properties. We adopted the following properties of this cluster after examination of values from different sources: the cluster’s metallicity ([Fe/H]=-0.46 dex) [24], half-light radius (1.02 arcmin) [24], distances (Rs = 11.3 kpc and Rgc = 4.2 kpc) [24] and reddening ($E(B-V) = -0.15$) [24].

However, in terms of variable stars, this cluster hasn’t been studied closely, and no LPV was detected in this cluster before. For instance, it was observed by Fourcade and Laborde and they found no variable stars [2]. Moreover, this GC is not as crowded as many of the other clusters and that might be the reason it has received less attention in terms of the search for variable stars. Many astronomers would choose to study GCs that are crowded with stars because statistically speaking they might be able to detect more variable stars and thus any conclusion that they would come up with will be based on a relatively large number of variable stars. Furthermore, GCs with metallicities larger than -0.8 dex are poorly studied because of observational difficulties in their location.

In 1996, Dr. Andrew Layden observed NGC 6496 cluster for 2 months with a 0.9m  

\[ Rc = \text{Distance from Sun (kiloparsecs)}, \quad Rgc = \text{Distance from Galactic center (kpc), assuming } R_0 = 8.0 \text{ kpc} \]
telescope and he was able to detect 5 LPVs, 3 short-period variable stars and a few other possible short-period variable stars. His study was not sufficient to determine accurate periods for LPVs because most of the LPVs have periods greater than 30 days so observing a LPV for 2 months is not sufficient to determine a reliable period.

Because of the variables he found and considering all the challenges and difficulties, we decided to conduct a search for LPVs in the metal-rich cluster NGC 6496 ([Fe/H]=−0.46 dex). NGC 6496 is open enough to permit photometry of stars all the way to the cluster center. This is a great advantage that will leave us with fewer errors and more reliable results. In addition, the cluster has a low reddening, $E(B-V)=-0.15$ [24].

In this paper we write about our work in detail regarding observing the cluster, processing the images, doing photometry, and searching for variable stars, and finally we show our results.
CHAPTER 2
OBSERVATION

2.1 Instruments

Images of NGC 6496 were taken with the Panchromatic Robotic Optical Monitoring and Polarimetry Telescope 4 (PROMPT 4) located at the Cerro Tololo Inter-American Observatory (CTIO) in the Chilean Andes (latitude-30:09:55, longitude-70:48:52) and owned by the University of North Carolina at Chapel Hill. Prompt 4 is a 0.41-meter diameter Ritchey-Chretien telescope [9] with a field of view of 10x10 arcminutes and a resolution of 0.59 arcseconds per pixel. PROMPT 4’s CCD camera has gain of 1.5 electrons/adu and readout noise of 13.5 electrons.

2.2 Imaging

We remotely\(^1\) observed the cluster from February 2009 until October 2010 during which we were able to image the cluster for 73 nights. Roughly, we took images every two weeks and we made sure that the airmass didn’t exceed 2.2 to limit the amount of atmospheric extinction and reddening. For our study, 10 nights were deleted from the data set due to bad seeing, bad weather, problems in tracking or other technical problems, leaving us with 63 useful nights. We divided our data into two seasons (see table 2.1). Season one contains observations between February 2009 and June 2009 and season two contains images between July 2009 and October 2010.

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Table 2.1: Different Seasons

\(^1\)We requested observation data online and we downloaded the images the night after.
On each night, 10 different images were taken with two different filters with three different exposure times. For each night, we requested 4 images in the V filter each with an exposure time of 80 seconds, 3 images in the I filter each with an exposure time of 40 seconds (IL) and 3 images in the same I filter each with an exposure time of 10 seconds (IS). This information is summarized in table 2.2.

<table>
<thead>
<tr>
<th>Filter</th>
<th>V</th>
<th>IL</th>
<th>IS</th>
</tr>
</thead>
<tbody>
<tr>
<td>Exposure(s)</td>
<td>80</td>
<td>40</td>
<td>10</td>
</tr>
<tr>
<td>Images</td>
<td>4</td>
<td>3</td>
<td>3</td>
</tr>
</tbody>
</table>

Table 2.2: Exposure time in different filters.

We requested images with the same filter (I) but with two different exposure times to make sure that we are able to detect stars that are faint and at the same time in order not to lose data for bright stars that can saturate the CCD. In other words, most probably stars that are very bright will be saturated in the IL (and V) exposures so it will be hard to do accurate photometry on them. On the other hand, it is very unlikely for these stars to be saturated in the IS exposure where the exposure time is only 10 seconds. In addition to that, stars that are faint might not be detected in the IS exposure but are more likely to appear with a longer exposure time (IL and V). We used two filters with three different exposure times because we wanted to compare the light curves of variable stars in the two filters with the three different exposure times. If the light curves behave in the same way then we can be more confident about our results. If the magnitude of the star was increasing and decreasing at the same time in all three exposures, then most likely this variation is due to a real change in the star’s magnitude. However, we might want to question any star that doesn’t behave similarly in the V, IL and IS exposures. Finally, the advantage of using two different filters (i.e. colors) is that it will allow us to draw the CMD of the cluster in order to look at where the variable stars fall on the diagram. This can help us know in which stage of evolution these stars are and it can tell us if these stars fall on the principal sequences of
the cluster, and therefore are likely members of the cluster. However, if the stars don’t fall on the principal sequences of the cluster, then most probably the stars are field stars.

For our observations, we requested the images to be taken in an interleaved order. This means that the second image should be taken in the I filter if the first was taken with the V filter and vice versa (V, IL, IS, V, IL, IS,...). The advantage of using this order over the sequential (V, V, V, V ..., IL, IL, IL, IS, IS,IS ...) order is that the cluster center will fall on a slightly different pixel every time an image is taken. In this case, if a star falls on a dead pixel in the first V image, it is highly unlikely that it will fall on the same dead pixel in the second V image. The same thing applies to the I filter. So it will be very unlikely that we lose any star due to any pixel deficiency. The reasoning for this will be discussed further in section 2.7.

2.3 Image Processing

In order to get more accurate and reliable measurements, image processing and calibration is essential. In general, three different sets of calibration images are needed for this process: bias, dark and flat frames. All of these frames are taken automatically every night by the PROMPT telescope system, and were simply downloaded from PROMPT’s telescope website (SKYNET). The images downloaded from SKYNET are called “raw” images and are referred to as “processed” after they are calibrated.

2.3.1 Bias Frame - B

Even when the shutter of the CCD camera is closed, it will still have a small intensity (bias) and the bias intensity usually varies across the CCD. The bias signal (or bias level) is uniform in every exposure (except for the read noise, which is overprinted on it) and it is the uniform component that we aim to remove [23]. We need to measure this bias signal and subtract it from our raw images. We do that by taking 0 second exposure images with
the shutter and dome closed. We then average these bias frame images to get one master bias frame image. Finally, the bias frame is subtracted from all the raw images to remove the baseline count in each pixel. PROMPT 4 takes 20 bias frames for each night and hence we averaged all these bias frames to de-bias the raw images. We average these bias frames to improve our signal-to-noise ratio (S/N). In addition, cosmic rays detection becomes easier when we use as many bias frames as we can. More details about this can be found in section 2.7.

2.3.2 Dark Frame - \((S)_d = (R)_d(t)_d\)

One can think about dark frames as bias frames but with non-zero exposure time. The CCD will still have a signal when the shutter is closed and this signal is proportional to the exposure time [10]. This intensity appears because of the random movement of electrons on the CCD chip. Unlike the bias frames, dark frames depend on the temperature of the CCD. Specifically, there is a logarithmic relation between the dark count rate and the temperature. In other words, “the dark count rate \((R_d)\) doubles by each change of a certain number of degrees” [10]. This is because electrons are less likely to move as the temperature goes down and they are more likely to move when the temperature rises. Thermal energy is proportional to the temperatures \((E \sim KT)\) so the electron movements generate higher biased intensity when the temperature is high because electrons would have enough thermal energy to overcome the binding energy of the lattice and become free [23]. This is one of the reasons CCD cameras need to be cooled. In fact, dark count rates don’t depend on temperatures only but on the exposure time as well. This can be summarized in equation 2.1 [11] and equation 2.2 [11]:

\[
R_d = R_{d0} \times 2^{(T - T_0) / \Delta T},
\]  

(2.1)
where $\Delta T$ is the doubling temperature, a property of the CCD, $R_{d0}$ and $T_0$ are the reference dark count rate and temperature respectively [11]. In other words, knowing the temperature of the CCD is sufficient to predict the dark count rate. Finally, the dark signal ($S_d$) can be calculated by multiplying the dark count rate by the exposure time:

$$S_d = R_d t_d$$  \hspace{1cm} (2.2)

From equation 2.2, one can realize that the longer the exposure time, the more “dark signal” we will have. Because our dark frame are of 80s exposure time and our raw frames are of different exposure times, we scaled the dark frames exposure times for each of our images. Just like bias frames, since the dark frame contains dark signals that we don’t need, it needs to be subtracted from the raw image. In order to improve the S/N of the dark frame, we need to take many dark frames and stack them together to get one “master dark frame” to subtract from our raw images. PROMPT 4 takes 20 dark frames for each night and hence we averaged all these frames to get our “master dark frame”.

### 2.3.3 Flat Fields - $(V)_sq$

One more frame is needed to fully calibrate a raw image: a flat field. Flat fields are used to correct many of the deficiencies in the process of CCD imaging. First of all, CCD chips are made up of very sensitive pixels, but not all the pixels will have the same sensitivity. In other words, each pixel will react slightly differently when exposed to the same light intensity [23] and we need to know what fraction of incoming photons are detected on each pixel.

In order to detect this pixel-to-pixel difference in sensitivity, the CCD camera should be exposed to an object that is uniformly bright. The most practical and reasonable source of light is the twilight sky. The sky should be clear from clouds when taking flat fields because the presence of clouds blocks the sky light unevenly. Exposing the CCD to such a uniform source of light should tell us how each pixel is reacting to the incoming photons.
Second of all, incoming light is usually reduced by the dust found on the filters. Taking flat fields can show us where the dust particles are located. Because dust appears in different locations on each filter, flat fields should be taken in each filter used. In order to improve the signal-to-noise (S/N) of the flat fields, we need to take many flat fields and stack them together to get one “master flat field” before dividing it into our raw images. PROMPT 4 takes 10 flat fields per night per filter and hence we averaged all these frames to get our “master flat field” in V and in I. The same flat fields were used for the IL and IS exposures.

The signal read on the CCD from striking photons is thus a combination of the actual signal from the photons \(R_0t_0\) multiplied by the pixel-to-pixel sensitivity map \(V_{sq}\) obtained from the flat fields, in addition to the bias signal \(B\) obtained from the bias frames and the dark signal \(S_d = R_dt_d\) obtained from the dark frames. Following Newberry, this can be illustrated in equation 2.3 [13]:

\[
S = R_0t_0V_{sq} + R_dt_d + B.
\] (2.3)

Hence, we should first subtract the bias and dark signal to be left with equation 2.4 [13]:

\[
S = R_0t_0V_{sq}.
\] (2.4)

Consequently, \(S\) should be divided by \(V_{sq}\) in order to get the real signal coming from the star \(R_0t_0\) [13]):

\[
R_0t_0 = \frac{S}{V_{sq}}.
\] (2.5)

<table>
<thead>
<tr>
<th>Image Type</th>
<th>Bias Frames</th>
<th>Dark Frames</th>
<th>Flat Fields</th>
</tr>
</thead>
<tbody>
<tr>
<td>Image/Night</td>
<td>20</td>
<td>20</td>
<td>10/filter</td>
</tr>
</tbody>
</table>

Table 2.3: Calibration Images.
2.4 Making Calibration Images

Since PROMPT 4 is a robotic telescope, all the raw images, bias frames, dark frames and flat fields were simply downloaded from PROMPT’s website\(^2\). We had to request our raw images individually every night. However, our raw images were not calibrated and processed. In order to get rid of all the CCD and instrumentation deficiencies discussed in the previous section, we used a program called the “Image Reduction and Analysis Facility” (IRAF) [17]. The first step is this process was to build our master bias frame, which we called the ZERO image. After that, we built our master dark frame, which we called the DARK image, and finally we built the master flat image for each filter. Since we used two different filters (V and I), two master flat images were built; we called them FLATV and FLATI. Note that these master images were built for every observing night because the dust on the filters and the pixel-to-pixel sensitivity may change from one night to another. Since we have three different sets of images (V, IL, IS), images with good seeing and with the same filter and exposure time were stacked to form a master image in each set for each night. Then, the BIAS and DARK images were subtracted from the raw images (V, IL, IS), FLATV was divided into the V image while FLATI was divided into the IL and IS images. All this work was done with IRAF.

2.4.1 ZERO Image

We combined the individual bias images into a single, averaged and cosmic-ray-rejected bias image (ZERO). To do so, parameters of the CCD camera that we are using (i.e. RD-NOISE and GAIN) needed to be set by the “epar” command in IRAF. After setting the parameters, the “imcombine” command was used to add and average all of the 20 individual bias images to produce our ZERO image, using a “sigma-clipping” process for bad-pixel rejection as explained in section 2.7.

2.4.2 DARK Image

In order to build our DARK image, we first subtracted our ZERO image from all the 20 dark frames. The “ccdproc” command in IRAF was used to do this job. “ccdproc” removes the bias level from each dark frame based on the ZERO image. After this step, the “imcombine” command was used again to average and reject bad pixels in all 20 individual dark images to produce our DARK image.

2.4.3 FLATV and FLATI Images

The ZERO and DARK images were subtracted from each of our flat images (10 in each filter, a total of 20). To do so, the “ccdproc” command was used again where this time it didn’t only remove the bias level on the individual flats, but it also scaled the DARK image to match the exposure time of each flat and it subtracted off the scaled DARK image to remove the dark current [14]. Then the “flatcombine” command was used to combine the individual flat images into a single, averaged and cosmic-ray-rejected image. The result of this process is two master flat images for each night.

Now that we have all of our calibration frames prepared, we need to use them to improve the quality of our raw cluster images. This takes us to the next step: calibrating and combining our raw images.

2.5 Calibrating the cluster raw images

The “ccdproc” command was used again to calibrate the cluster raw images. “ccdproc” removed the “bias level” by subtracting the ZERO image from each raw individual image. It also scaled the DARK image to match the exposure time of the raw images in order to subtract off the right amount of dark current. Finally, it divided the raw images by FLATV or FLATI (depending on the raw image). Now all the individual raw images are calibrated and are ready to be stacked.
2.6 Combining raw images

We want three final raw images for each observing night, one in V, one in IL and one in IS. In other words, the four individual cluster images in the V filter need to be stacked to give us one “Master V Image”. Same thing applies to the IL and IS images, keeping in mind that three individual images were taken in these filters. The CCD camera’s position and the telescope’s alignment change with time so it is almost impossible for the cluster center to fall on the exact same pixel in every exposure. However, the raw images need to be aligned in order to be stacked properly. Consequently, the first step in combining the raw images was done manually by choosing a reference star and the cluster center in each image. The X and Y coordinates of the reference star and the cluster center were measured and saved for all the images. Some of the images were rotated by 180 degrees and so we had to flip them so that they matched the orientation of the other nights. A simple FORTRAN program\(^3\) was used to calculate the X and Y shifts of the reference star between each image and a “reference” image of our choice. The reference image was chosen to have good seeing and a small FWHM. The reference star in the IL reference image is shown in figure 2.1.

![Reference Star Image](image)

Figure 2.1: The first and third images represent the same cluster image. However, it is easier to identify the reference star in the third image. The second image represents the radial profile of the selected reference star.

\(^3\)Written by Dr. Andrew Layden
Ideally, 4 images should be stacked to form our master V image, 3 images should be stacked to form the IL image and same for the IS image. Although stacking images in the same filter and with the same exposure time improves the S/N and helps us detecting fainter stars, some of the images were not used in the stacking simply because of their bad quality. To ensure that the stacking process will give us the best results, one should take a look at the background level of the images before stacking them. Seeing varies from night to night due to various atmospheric conditions. The “imstat” command in IRAF gives us information about the background level. Using the “imarith” command, we made all the images to have similar background levels (less than 10% difference). Finally, the stacking task was done again with the “imcombine” command in IRAF. The result of this was three final images in each night, one in the V filter and two in the I filter (IL and IS). The power in stacking (adding) images is that it produces images with higher S/N and with better resolution because it increases the S/N ratio, which of course can help us detect faint stars. Noise (N) adds quadratically while signal (S) simply adds linearly [12]. This can be represented in equations 2.6 [12] and 2.7 [12]:

\[ N^2 = N_1^2 + N_2^2 + ... \] (2.6)

\[ S = S_1 + S_2 + ... \] (2.7)

And dividing the signal by the noise gives us equation 2.8:

\[ S/N = \frac{(S/N)_1}{\sqrt{1 + (N_2/N_1)^2}} + \frac{(S/N)_2}{\sqrt{1 + (N_1/N_2)^2}} + ... \] (2.8)

However, in our case, the added images had the same exposure time and they were taken under similar conditions so we expect them to have the same signal and noise [12]:

4This could be due to clouds passing by during the exposure, problem in tracking, etc.
5reasons like clouds, change in airmass, etc.
\[ N_1 = N_2 = N_3 = ... \]  

\[ S_1 = S_2 = S_3 = ... \]  

Consequently, equation 2.8 becomes [12]:

\[ S/N = \frac{(S/N)_1 + (S/N)_2}{\sqrt{1 + 1 + . . .}} \]  

\[ S/N = \sqrt{n}(S/N)_1 \]  

So adding “n” images with the same exposure time and conditions increases the signal by the square root of n.

### 2.7 Bad Pixel Rejection

Bad pixels are pixels whose measurements can’t be relied on. These pixels may be struck by cosmic rays, which arrive at random times and places on the CCD. These pixels might also not be functioning properly, returning either no signal or a signal that is not proportional to the number of photons that fell on them (“bad pixels”) [15]. Other pixels may have high dark currents (they produce a large number of thermal electrons per second unrelated to the flux of photons falling on them) that can’t be corrected properly by our usual dark correction procedure (“hot pixels”). Fortunately, the number of such discrepant pixels tends to be small, though the number of cosmic ray hits and the severity of hot pixels do increase with the exposure time. Taking many images for our object and averaging them is a useful way to reduce the noise [15]. On the other hand, we use a process called “sigma-clipping” to reject cosmic rays effects. This process averages out the intensity and the standard deviation at
each pixel and rejects the outliers and recalculates the intensity and the standard deviation. In this way, we make sure that we are getting rid of any intensity that was taken for a cosmic ray. The “sigma-clipping” process can be useful in rejecting bad and hot pixels if one moves the telescope few arcsec between exposures [15]. This will insure that a different star or sky object will fall on the bad pixel in each exposure so the “sigma-clipping” process will reject them [15]. In all cases, finding the average and the standard deviation for each pixel is crucial when it comes to rejecting bad pixels. Hence, taking many frames and averaging them makes the bad pixel rejection more reliable because the intensities of these bad pixels will be compared to an averaged intensity that came from many frames (see Figure 2.2).
Figure 2.2: Increasing the numbers of frames averaged out increases the S/N and rejects most of the bad pixels [15]
CHAPTER 3
PHOTOMETRY

3.1 ISIS Photometry

The first program used for photometry is called ISIS (by Alard and Lupton [18]). Unlike other photometry programs (e.g. DAOPHOT), ISIS doesn’t use the aperture/PSF fitting photometry method. This program works on a simple subtraction method. In order to do so, all frames need to be aligned for ISIS to subtract the frames correctly. We had to trim many parts of the frames from different nights so that all the images are aligned so that we can take the mathematical intersection of all the images to apply the subtraction method. Before applying the subtraction method, a master image was created. ISIS takes into consideration that the seeing varies from image to image. Hence, the master image gets blurred to match the seeing of each frame so that the subtraction method works well. If the same star on two images has the same brightness level, then the subtracted image shouldn’t contain any signal. However if a star has different brightness levels in two frames then the subtracted image should contain a signal that would represent the difference in the brightness level. These are our prospective variable stars. Theoretically, the subtracted images should include signal from variable stars and the constant stars should cancel out. The subtraction method is illustrated in Figure 3.1. ISIS counts the number of photons (flux) of each variable star and gives an estimate of the brightness level. It has many great advantages. This subtraction method works well in the crowded regions (near the cluster center) and it is relatively easy and fast. However, this method doesn’t work as well near the edges of the frames.

Because we have frames extended over 600 days, we expect the frames to have different centers and orientations. In each frame, we determined the horizontal and vertical position of a reference star of our choice. As a second point, we also estimated the cluster center in each frame in order to help aligning and trimming all the frames. After getting the coordinates
Figure 3.1: Subtracting two frames, illustrating stars with the same and different intensities.

needed to align all the frames to the same coordinate system, we chose the frame with the best seeing (see Figure 3.2 and Figure 3.3) and we called it the reference frame. We trimmed the frames to make sure that they all have the same number of pixels and that every star falls on the same pixel in every frame.

Each frame has a different point spread function (PSF) and seeing mainly because these frames were taken on different nights. Frames from different nights were taken in different weather conditions and some frames had tracking problems. Which caused the frames to have different PSFs and seeing. The seeing was measured by the FWHM of the radial profile of one of the uncrowded bright stars. Having different PSFs causes problems in matching the seeing of different frames. Alard and Lupton adopted a method to find the least-squares solution of equation 3.1:

$$Ref(x, y) \otimes Kernel(u, v) = I(x, y)$$  \hspace{1cm} \text{(3.1)}
where $Ref$ is the reference image, $I$ is the image to be aligned and the symbol $\otimes$ denotes convolution. However, because each frame has a different background, equation 3.1 should be corrected to become equation 3.2:

$$\text{Ref}(x, y) \otimes \text{Kernel}(u, v) = I(x, y) + \text{bg}(x, y)\quad (3.2)$$

where $\text{bg}(x, y)$ represents the differential background variation [18].

After choosing the reference frame and estimating the coordinates of the reference star and the cluster center, we ran the “interp.csh” command in ISIS. Interp.csh solves for and removes any shift and small rotations between each frame and the reference frame. After
Figure 3.3: The surface plot and the radial profile of the reference star in Figure 3.2.

this step all the frames should be aligned very precisely to the reference image in order for
the photometry process to work well. In order to get better subtracted frames, we chose the
best (around 10%) frames and stacked these frames together using the “ref.csh” command.
We got one frame with better S/N and with fewer cosmic rays effects (since we are averaging
out many frames) and we called it “ref.fits”.

At this point, we started the subtraction process with the ISIS command “subtract.csh”.
ISIS does pixel to pixel subtraction between “ref.fits” and each frame. Theoretically, only
the variable stars should be seen on the subtracted frame (see Figure 3.1). Most of the time,
the variability of the stars is small and it would be difficult to see it on the subtracted image
as it can be easily mistaken for noise. In order to overcome this problem, we need to stack
all the subtracted images.

Stacking all of the subtracted images can be done with the “detect.csh” ISIS command,
which adds all the signals. The stacked image is called “var.fits” (see Figure 3.4). Any
signal on this image is a possible variable star. ISIS can automatically locate these signals
and can give us their positions with the “find.csh” command. Before running the “find.csh”
command, it is important to set a threshold intensity where any signal with intensity less
than our threshold intensity will be ignored (although it could be a possible variable). Our
threshold intensities were chosen to be 0.5 for IL, IS and V. The threshold intensity was
chosen by examining all the signals on the subtracted image and estimating that any star with an intensity less than 0.5 is probably not a variable star. Finally, running the “phot.csh” command gets the fluxes of the possible variable stars detected with “detect.csh” and saves them in a text file. At this point, we have all possible variable stars with their flux values over time and hence we are ready to draw the light curves and calculate periods.

Figure 3.4: Var.fits in the IL filter. All of the white circles are variable stars candidates.

However, one cannot completely rely on ISIS to detect variable stars for a few different reasons. First of all, ISIS calculates the fluxes of stars instead of magnitudes and usually stars are so crowded and blended with other stars on ref.fits that we can’t rely on determining its magnitudes from their fluxes. So even if we were able to detect variable stars and to get
their periods, we won’t be able to rely on the locations of the stars on the CMD that was based on fluxes from ISIS.

In addition, as mentioned before, all the frames need to be aligned for ISIS to subtract the frames correctly. In order to do that, we had to trim many parts of the frames from different nights so that they would all have the same number of pixels. The disadvantage of this step is that we will be losing stars near the edges of the frames. Moreover, although the ISIS subtraction method is powerful in crowded regions (i.e. the cluster center), it is not so effective at the edges. This is because the frames’ alignment will be best in the center and it will get worse as we go further out. Not only that, the focus of each frame changes slightly from the center of the frame to the edges. The results of these disadvantages appear in the subtracted image “var.fits”. The locations and radial profiles of two stars (one near the center and one near the edge) are shown in Figures 3.5 and 3.6. Notice that the radial profile of the star near the center is well fitted but the radial profile fitting is poor for the star next to the edge of the frame. We used such plots to set our threshold intensity.

3.2 DAOPHOT Photometry

After using ISIS, a second method was adopted to do photometry, using a collection of softwares by Stetson: DAOHOT II, ALLSTAR, ALLFRAME and DAOMASTER. Although ISIS works better in crowded regions (especially in the cluster center), the advantage of using DAOPHOT is that it measures the brightness of every star on every frame, and then allows us to search for stars that vary in brightness [23]. In addition, it will enable us to draw a color magnitude diagram for our cluster since unlike ISIS, DAOPHOT calculates the actual magnitude of each star. Finally, detecting the same variable star using two different methods will make us more confident about the results of these stars.

Several tasks, packages and scripts were used in this long photometry process. The first task used was “FIND”. The FIND task is used to find objects above an assigned threshold.
Moreover, it distinguishes legitimate stellar images from random noise peaks and cosmic rays. It also recognizes when an extended object consists of two or more overlapping stellar images. The properties of our CCD and the FWHM of a reference star in each frame were provided to DAOPHOT [19]. FIND task examines each frame pixel by pixel and it looks for intensities above a threshold intensity of our choice. If it detects a bright spot, it fits the star with a Gaussian profile that is proportional to the brightness of the star. On the other hand, if the brightness level in a pixel was less than our threshold value, the central height of the Gaussian will be near zero. This gives us a clear picture of where the stars
are approximately located and what their brightness levels are. Now that the locations of all the stars on all the nights are found, aperture photometry is done using the “PHOT” task in DAOPHOT. “PHOT” goes back to each star found in the “FIND” command and adds up the pixel counts within a circle centered on each star and subtracts the average sky brightness. The reason we subtract the sky brightness is that some of the photons detected on the CCD from each star are coming from other sources like scattered light from other stars, the moon etc.

The question that the “PHOT” task and the aperture photometry in general try to answer is the following: “What would the brightness level be in each star location if the star was not there?” Answering this question gives us an estimate of the sky background which needs to be subtracted from the star’s image. Visually (see Figure 3.7), “PHOT” locates the outer sky region, the inner sky region and the radius of each star. It counts the photons/brightness in the outer sky region, averages it and subtracts it from the inner region. This gives us an estimate of the actual brightness of the star. Note that the chosen radii are proportional to the intensity of every star.
Figure 3.7: The “PHOT” task locates an inner (white) and outer (blue) regions around the star and subtracts them to get a more reliable brightness level for the star.

Now that the location and brightness of each star has been calculated, point spread functions (PSFs) of all the stars should be scaled to one averaged PSF in order for magnitudes to be calculated. Before building the PSF model, a number of bright isolated stars that are not saturated should be selected and averaged to give our PSF model. The “PICK” task in DAOPHOT makes this step easy as it automatically selects stars with such properties. Because the population of stars decreases as we move away from the center. It is important for PICK to pick stars away from the center to avoid any crowded star so that we can build a more reliable PSF profile. The PSF stars selected by DAOPHOT on one of the nights in the IL filter are shown in Figure 3.8. After selecting the PSF stars, the “PSF” command was used to combine all the PSF stars’ profiles into a PSF model. Although we requested 160 PSF stars, an average of around 145 PSF were actually used because DAOPHOT rejects any star that falls on a bad pixel or a star that has a lot of neighboring stars. All of the PSFs were scaled to match each star’s intensity in the image. This will help DAOPHOT later to calculate the magnitude of these stars. This step is done using the “ALLSTAR” program, which fits the PSF model to all of the stars in each image. In order to do that,
Figure 3.8: Around 160 PSF stars were selected using the “PICK” task. Notice that stars in the center were neglected and that the brightest unsaturated and isolated stars were selected. Also notice that the selected PSF stars are distributed all around the cluster.

ALLSTAR uses the position (X and Y) and the magnitude of each star that we found using the previous commands. This step is done over many iterations (see Figure 3.11). A new file is created containing a more accurate measurement of the stars’ positions and magnitudes as well as the number of iterations it took ALLSTAR to fit the stars to the PSF model. At this step, we have an output file with good estimates of the stars’ magnitudes and positions and the sky brightness. However, the PSF model was affected by stars that are close to the PSF stars chosen. Overcoming this problem can be done using the “SUB” task on all of the images.

“SUB” generates a new PSF model after subtracting neighboring stars from the PSF
stars. It does this by generating a file containing the positions and magnitudes of all of the stars near the PSF stars. Then it simply subtracts this list of stars from the corresponding raw image, leaving us with a subtracted image containing the PSF stars but with fewer neighboring stars (see Figure 3.9). Now a more reliable and improved PSF can be found using the “PSF” task again. Moreover, “ALLSTAR” is run again using the improved PSF model and a new improved PSF-fitted photometry file is generated.

Figure 3.9: Part of the cluster image after applying the “SUB” task. Red circled stars represent PSF neighbor stars while the green circles represent the PSF stars. Notice that most of the stars near the PSF stars have disappeared, allowing a better PSF model to be built.

Finally, the “FIND” and “PHOT” tasks were performed to give us a new list of stars. We ran FIND/PHOT with a lower detection threshold this time on the subtracted image to find entirely new stars, often fainter and in the edges of brighter stars. The two lists are then combined to give us one final list of stars with their corresponding properties (positions, magnitudes, sky brightness, errors, etc.). All of these steps were run on each night and the results, comments and details about each run were saved in a separate file for each night (called NIGHTS’DATE.log).

Many problems occurred while running all the tasks above for all of our images. For
instance, some of the PSF stars that were selected were overlapped by other neighboring stars. This led to a bad PSF model and eventually it was hard to fit the PSF model to all of the stars. We were able to identify these problems by looking at the “NIGHT'SDATE.log” file to see whether if “CHI” was converging or not. Part of a “.log” file is illustrated in Figure 3.10.

![Part of a "log" file](image)

Figure 3.10: Part of a “.log” file. Notice that CHI converged from 0.0318 to 0.0176. The profile errors are the selected PSF stars and the errors on them. We know that this night worked fine because CHI converged and also because all of the PSF stars have small errors.

As mentioned in the chapter 2, we divided our nights into two separate sets. The division was not necessary for all of the preceding steps but it was necessary for using “DAOMASTER”. The main purpose for using this program is to match the same stars in different frames. The “FIND” task used earlier worked on each frame separately and it

---

1CHI is a measure made by DAOPHOT of how well a star’s profile matched the PSF
assigned ID numbers for stars differently on each frame. For instance, the same star can have an ID of 20 on the first frame and an ID of 45 on the next frame. However, it is crucial to know the ID number of each star on each image so that we can study its magnitude variation. This matching process was done through many iterations. The first step was to get an initial estimate of the position relationship between the master image and each of the subsequent images for each season in each filter. To do this, we provided the program with the X and Y coordinates of a reference star in each image. Then, what we call a MONTAGE image was built. We wanted to get a uniform reference star list with as many stars as possible (especially faint stars). We selected the best images from the IS exposure and combined them to create a Montage image for our IS images. Similarly, we combined the best IL and V images to create a common montage image for the IL and V exposures. After creating our two montage images for each season, the DAOPHOT and ALLSTAR tasks were performed on these images to create the star lists.

After DAOPHOT/ALLSTAR reductions, ALLFRAME was used for an additional and more accurate measurement. Each image is re-processed using the same list of star magnitudes and positions. This may not have much effect on the best images but it will have a great effect on images taken on cloudy nights because only few stars would be detected with the original DAOPHOT/ALLSTAR processing on these nights (or on images with very bad seeing). Here we feed the known positions of the stars to the image, and it will do its best to fit them and get magnitudes (though the errors may be large). It is very important for variable star work to attribute the light to a consistent set of objects (e.g. same amount of light attributed to fewer stars means each star looks brighter) [23]. DAOMASTER was performed on all the nights using the more accurate magnitude and position values calculated by ALLFRAME. The result of this was files containing the average magnitude of each star with the corresponding error and the magnitude of each star on each night with its corresponding error.
Figure 3.11: The number of iterations taken for all of the stars to converge. Most of the stars converged on this night, which means that the PSF model was well selected and well fitted to all of the stars.
CHAPTER 4
CALIBRATION

At this point, all the magnitudes that were detected are instrumental magnitudes (v and i). In other words, DAOPHOT calculated the magnitude of each star using its own arbitrary zero point which means that all of the magnitudes at this step are systematically off their standard values. In order to find this offset and calibrate these magnitudes to get the standard magnitude of each star, we used three different studies that were done on NGC 6496. What we need from other studies is a set of standard magnitudes of stars in and around our cluster. Comparing the standard values (V and I) with our values and fitting the relationship between these two will allow us to correct our instrumental magnitudes.

The first data used for the calibration process were from Stetson [20]. Stetson’s website provided us with a list of 34 stars around our cluster with their magnitudes in V, R and I filters, as well as his image of the cluster. Visually, we located the coordinates of the 34 stars on our image and we designed a table containing the ID number of each star and the V and I magnitudes from both studies (Stetson’s and ours). The locations of these stars are shown in Figure 4.1.

Then, we plotted the relation between \( (V - I) \) and \( (v - V) \) where \( v \) corresponds to our instrumental magnitude and \( V \) and \( I \) represent the standard magnitudes provided by Stetson for each star. This is illustrated in Figure 4.2. The fitting of the this plot will be performed next but one can see that the shape of this plot doesn’t require a high order fit (see Figure 4.2). The curfit IRAF command was used to fit each of the following plots. The corresponding plot of \( (V - I) \) and \( (i - I) \) is shown in Figure 4.3.

The second study that we used to calibrate our magnitudes was that of Armandroff [21]. Armandroff provided a list of 285 stars with standard V and (V-I) magnitudes. Visually, we located 30 of his standard stars on our image and we used a FORTRAN program called FITROT [23] which calculated the rotation, the offset and the scale difference between our
and Armandroff’s image. The list of 30 stars was used to set up FITROT. The entire list of 285 stars was used for calibration after we used FITROT. Then, the program used these values to convert the whole Armandroff list to match the X,Y coordinates of our cluster image. The locations of his stars on our image are shown in Figure 4.4.

Again, we plotted the relation between \((V - I)\) and \((v - V)\) as well as \((V - I)\) and \((i - I)\), where \(v\) and \(i\) correspond to our instrumental magnitudes while \(V\) and \(I\) represents the standard magnitudes provided by Armandroff for each star. These plots are shown in Figure 4.5 and Figure 4.6. The general equation which represents the linear fitting is of the form \(Y = A + BX\) where \(Y\) represents the difference between our magnitudes and Armandroff’s (or Stetson’s) magnitudes while \(X\) represents the \((V - I)\) magnitudes from Armandroff (or Stetson). Finally, \(A\) and \(B\) represent the zeroth and first order coefficients, respectively (see equations 4.1 and 4.2) [23].

The fitting was done 6 times. Specifically, we compared Stetson’s magnitudes with our \(v\), \(i\) and \(il\) observations and we did the same thing for Armandroff’s data. We used zero and first order fitting to fit these graphs and the coefficients of the fittings are summarized in table 4.1. Notice that a zero order fitting was sufficient to fit IL to Armandroff’s data and \(V\) to Stetson’s data. We used the zero order fit for these two fits because the first order fit yielded errors in the coefficients greater than the \(A\) and \(B\) coefficients themselves.

\[
v - V = A_v + B_v(V - I) \tag{4.1}
\]

\[
i - I = A_i + B_i(V - I) \tag{4.2}
\]

Subtracting equation 4.2 from equation 4.1 gives:

\[
(v - i) - (V - I) = (A_v - A_i) + (B_v - B_i)(V - I). \tag{4.3}
\]
Solving equation 4.3 for \((V - I)\) gives:

\[
(V - I) = \frac{(v - i) - (A_v - A_i)}{(1 + B_v - B_i)}. \tag{4.4}
\]

Substituting the last expression for \((V - I)\) in equations 4.1 and 4.2 gives us the calibrated magnitude of each star in each filter (equations 4.5 and 4.6) [23]):

\[
V = v - A_v - B_v(V - I) \tag{4.5}
\]

\[
I = i - A_i - B_i(V - I) \tag{4.6}
\]

<table>
<thead>
<tr>
<th>Data</th>
<th>Filter</th>
<th>A</th>
<th>Aerr</th>
<th>B</th>
<th>Berr</th>
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<td>0.0106</td>
</tr>
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<td>0.0</td>
<td>0</td>
</tr>
<tr>
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<td>-0.0042</td>
<td>0.001</td>
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<td>0.0029</td>
<td>0</td>
<td>0</td>
</tr>
<tr>
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</tr>
<tr>
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<td>1.9100</td>
<td>0.0051</td>
<td>-0.0950</td>
<td>0.0082</td>
</tr>
</tbody>
</table>

Table 4.1: Coefficients of the Calibrations

Having these coefficients, we are ready to transform our instrumental magnitudes into standard magnitudes. To do so, a Fortran program \((TRANFORMvbri.f)\) [23] was used. The program asks for a list of stars and their corresponding magnitudes and errors. It then asks for the coefficients that we want to use (table 4.1) to make the conversion. Because we have two different instrumental I magnitudes (IS and IL), we made sure that the program used the “I” magnitude that had less error; from DAOPHOT. In this way, we can be more confident that our I standard magnitudes have small errors. After this, it applies the conversion to each star and gives us a new list of the same stars with their standard magnitudes that we are going to use to analyze the cluster’s CMD and variable stars. The result of this step was two lists containing the I and V magnitudes of all the stars in both filters based on
Armandroff’s and Stetson’s data.

In order to see if the whole calibration process was successful, we compared our final standard magnitudes with Stetson’s and Armandroff’s magnitudes in each filter. In addition, we examined the spatial dependence of the magnitudes (see Figure 4.7 and Figure 4.8). The plots show that the calibration process was successful because the difference of magnitudes in the V and I filters are centered at zero. In addition, there is no spatial dependence of the magnitudes because the difference of magnitudes was centered at zero regardless of the position of the star on the CCD. This is true for both data sets used.

4.1 Choosing the best calibration coefficients

In order to check our magnitudes further, we compared our standard magnitudes using Armandroff’s data with those using Stetson’s data to see if there is a dramatic difference between the two. Both can be seen in Figure 4.9. We can see that the I and V standard magnitudes from the two sets differ by around 0.05 mag.

Because we have two sets of standard stars and because Figure 4.9 shows that our final standard magnitudes are dependent on the calibration we are using, we had to decide which set to use. In order to choose, we found that the two sets have 17 stars in common and hence we decided to compare the magnitude of these stars. Figure 4.10 shows that the standard magnitude of the stars in V and I filters from the two sets are different and that there is a slight spatial dependence of the magnitudes.

At this point, the best way to decide which data to use for calibration was to compare with a third source. The third data set used for comparison was published by Richtler et al. [22]. Richtler et al. provided a list of around 200 stars in NGC 6496 with their corresponding standard V magnitudes. Comparing the V magnitude of the common stars between Stetson and Richtler et al. and between Armandroff and Richtler et al. is illustrated in Figure 4.11. One can see that the \((V_{st} - V_{rich})\) is on the order of \(-0.1\) while \((V_{arm} - V_{rich})\) is on
the order of $-0.2$. This shows that the Richler et al. data are about 0.1 mag fainter than Stetson’s, while Armandroffs are about 0.1 brighter than Stetson’s. For example, a star measured by Stetson to have $V = 10.0$ mag would be listed as $V = 10.1$ by Richtler et al. and 9.9 by Armandroff [23].

From Figures 4.10 and 4.11, the following relationships between the magnitudes were found and are illustrated in Figure 4.12.

\[
V_{st} - V_{rich} \approx -0.1
\]  \hspace{1cm} (4.7)

Solving for $V_{rich}$:

\[
V_{rich} \approx V_{st} + 0.1
\]  \hspace{1cm} (4.8)

\[
V_{arm} - V_{rich} \approx -0.2
\]  \hspace{1cm} (4.9)

Solving for $V_{rich}$:

\[
V_{rich} \approx V_{arm} + 0.2
\]  \hspace{1cm} (4.10)

\[
V_{st} - V_{arm} \approx 0.1
\]  \hspace{1cm} (4.11)

Solving for $V_{arm}$

\[
V_{arm} \approx V_{st} - 0.1.
\]  \hspace{1cm} (4.12)

From these relationships between the magnitudes, we decided to choose Stetson’s data for our final photometric calibration. Our choice was based on the fact that the magnitudes provided by Stetson fall between Armandroff’s and Richtler et al. magnitudes. Moreover,
Stetson used a more advanced CCD camera in his study as well as more advanced and improved DAOPHOT versions.

4.2 Calibrating the Variables Stars

Calibrating the variable stars is not as simple as calibrating regular stars. When we calibrated regular stars, we assumed that their magnitudes were fixed and not changing, which allowed us to use the calibration constants calculated in the previous section. However, when it comes to variable stars, there is no fixed magnitude that we can use for the calibration. One can use the mean magnitude of each variable star as the fixed magnitude. However, we decided to choose a more accurate and reliable method by comparing the magnitude of each variable star on each night with the magnitudes of up to 10 neighboring stars (comparison stars). So the calibration of each variable star was done separately on each night using the calibration constants derived by comparing our data with Stetson’s data. For a more precise and accurate calibration, we chose the comparison stars to fall within a close distance from the variable star to correct for any variation in the zeropoint with X,Y position from night to night (due to passing clouds, etc.) [23]. In addition, we chose the comparison stars to have colors close to the colors of the variable stars. To find comparison stars, we used a Fortran program (compselect.e) [23] which asks for the color range of our variable star and provide us with a list of comparison stars with colors close to the color of our variable star while making sure that the comparison stars are not variable stars themselves (Figure 4.14).

Mathematically, the first part of calibrating the variable stars is the same as calibrating regular stars (Equations 4.5 and 4.6) [23].

Now we derive the corrected magnitude of each variable star on each night in each filter using Equations 4.13, 4.14 and 4.14. These equations will use the instantaneous color of the variable stars on each night using Equations 4.5 and 4.6 [23].
\[ I = i - A_i - B_i(V - I) \]
\[ V_v = (v_v - v_c) - V_c - B_v((V - I)_v - (V - I)_c) \]  
(4.13)

\[ I_{L_v} = (i_{L_v} - i_{L_c}) - I_{L_c} - B_{iL}((V - I)_v - (V - I)_c) \]
\[ I = i - A_i - B_i(V - I) \]  
(4.14)

\[ I_{S_v} = (i_{S_v} - i_{S_c}) - I_{S_c} - B_{is}((V - I)_v - (V - I)_c) \]  
(4.15)

where \( V_v, I_{L_v}, I_{S_v} \) are the final corrected magnitudes for each variable star on each night in each filter while \( v_v, i_{L_v}, i_{S_v} \) are the instrumental magnitudes of the variable stars from DAOPHOT/ALLFRAME and \( V_c, I_{L_c}, I_{S_c} \) are the standard magnitudes. Applying these formulas using several comparison stars for the same variable star on the same night yields different possible magnitudes. We chose to use the median of these magnitudes to avoid any outliers and to reduce errors.

We also compute the standard deviation and standard error of the mean \( (\text{err} = \sigma/\sqrt{N}) \) (SEM); we use the SEM as the uncertainty in the median magnitude. It gives us an external estimate of the magnitude uncertainty based on real scatter from one star to the next, rather than the internal error based on the error estimates generated by DAOPHOT/ALLFRAME [23].

In this way, we obtained calibrated V and I magnitudes for each variable star on each of the 63 nights.
Figure 4.1: Location of Stetson’s standard stars on our cluster image.
Figure 4.2: Comparing the magnitudes of the 34 stars provided by Stetson with our instrumental magnitudes in the V filter and a linear fit.

Figure 4.3: Comparing the magnitudes of the 34 stars provided by Stetson with our instrumental magnitudes in the I filter.
Figure 4.4: Location of 285 standard stars that were used to get the calibration coefficients for our instrumental magnitudes from Armandroff’s data.
Figure 4.5: Comparing the magnitudes of around 250 stars provided by Armandroff with our instrumental magnitudes in the V filter and a linear fit.

Figure 4.6: Comparing the magnitudes of around 250 stars provided by Armandroff with our instrumental magnitudes in the I filter and a linear fit.
Figure 4.7: Comparing our final standard magnitudes with the Stetson’s data and examining the spatial dependence of the magnitudes.

Figure 4.8: Comparing our final standard magnitudes with Armandroff’s data and examining the spatial dependence of the magnitudes.
Figure 4.9: $V_{arm}$ and $I_{arm}$ are our standard magnitudes using the transformation coefficients calculated from Armandroff’s data set. $V_{st}$ and $I_{st}$ are our standard magnitudes using the transformation coefficients calculated from Stetson’s data set. One can see that the two sets of standard magnitudes do not agree.
Figure 4.10: Plotting the magnitudes of the 17 common stars between the two sets of standard stars used for our calibration shows that the two sets disagree on the magnitude of the stars.

Figure 4.11: Difference in V magnitudes between the three data sets used for calibration.
Comparing dV magnitudes

Figure 4.12: Difference in V magnitudes between the three data sets used for calibration. The blue open triangles represent the difference between Stetson’s and Armandroff’s magnitudes (+0.1). The red filled triangles represent the difference between Stetson’s and Richtler et al.’s magnitudes (-0.1). The black stars represent the difference between Armandroff’s and Richtler et al.’s magnitudes (-0.2).
Figure 4.13: Red triangles represent our possible variable stars and green triangles represent our comparison stars. Notice that there are comparison stars with colors relatively close to the colors of the variable stars. Not all of the stars that have colors close to the variable stars were chosen to be comparison stars mainly because they are close to the center (crowded region) where photometric errors tend to be larger.
Figure 4.14: Red circles represent possible variable stars and green circles represent comparison stars. Notice that around each variable star we can find 6 to 10 comparison stars with relatively similar magnitudes.
CHAPTER 5

COLOR MAGNITUDE DIAGRAM (CMD)

We plotted the color magnitude diagram of NGC 6496 for several reasons. It can tell us if our photometry went well from the shape of the plot. Since NGC 6496 is a globular cluster, we expect a CMD with an obvious red giant branch (RGB) and a horizontal branch (HB). A thin RGB will tell us that our photometry is good (small scatter), though a thick RGB could also be due to a range of reddening values due to variations in dust across the face of the cluster. This plot can also help us classifying our variable stars graphically. For instance, if one of the variable stars is red in color (i.e. large (V-I)) and if the star falls on the RGB and the period is more than 30 days, then most probably this is an LPV. We have to question any star that we think is an LPV that doesn’t fall on the RGB. In addition, the CMD can give us a hint about the membership of the detected stars. For instance, a variable star that is off the cluster’s RGB and HB is probably a field star. However, it doesn’t automatically mean that the star is a member of the cluster if it is located on the cluster sequences. The CMD of all of our detected stars is found in Figure 5.1. Notice that the star marked with a star at the top left is off the cluster CMD although it is less than 1 arcminute from the cluster center.

We plotted the CMD after calibrating all of the stars (including the variable stars) using the calibration method adopted by comparing our results with Stetson’s, shown in Figure 5.2. The other calibration method we adopted for variable stars was used to plot the light curves, not the CMD. The first part ((a) All Stars) of Figure 5.2 shows all of the stars we detected on our CCD field of view. Many of these stars are field stars and stars with large errors. Because photometric errors of stars usually increase in the crowded regions, we decided to plot the positions of the crowded stars on our CMD ((b) RjRin). Moreover, most of the field stars on our image must be located far from the center (since our CCD covers a
Figure 5.1: CMD of NGC 6496 after calibrating all the stars. One of the bright stars that is less than 1 arcminute from the cluster center is probably a field star because it is shifted from the CMD and doesn’t belong to the diagram. This star is marked with a black star.

larger area than the cluster itself\(^1\). In order to estimate the area of the cluster around the center, we calculated the tidal radius of the cluster to be \( R_t = 5 \) arcmin using the half-light radius\(^2\) of the cluster provided by Harris [24]. We also plotted all of the stars that are far from the center ((d) \( R > R_{\text{out}} \)). Finally, the CMD of the stars that are neither far from the center nor in the crowded region is the plot that we will consider as our improved CMD for now ((c) \( R_{\text{in}} \leq R \leq R_{\text{out}} \)). We chose the values of \( R_{\text{in}} \) and \( R_{\text{out}} \) to be 0.29 arcmin and 4.74 arcmin, respectively. In general, we chose \( R_{\text{in}} \) and \( R_{\text{out}} \) by trial and error. Since our cluster

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\(^1\)which is represented by the tidal radius

\(^2\)The half-light radius is the radius within which half of the cluster’s light is contained.
is not too crowded in the center, we started with very small Rin and we plotted figures similar to Figure 5.2. While increasing the value of Rin by small increments, we noticed that many of the stars that are off the main parts of the CMD (i.e. stars that don’t belong to the HB, RGB etc.) were being subtracted from our improved CMD. We tried Rin greater than 0.29 arcmin and we noticed that we were losing some of the stars on the RGB and HB and that are probably members of the cluster (based on their position on the CMD diagram and on their magnitude errors). We adopted the same technique to choose Rout. We started with Rout=6 arcmin and we plotted figures similar to Figure 5.2 while decreasing the value of Rout. We noticed that the main parts of the CMD (RGB, HB etc.) were getting more obvious and clear as we decreased Rout. We also noticed that many of the stars that were off the cluster sequence were getting subtracted. By trial and error and by observing the CMD plots we chose Rout to be 4.74 arcmin. Statistically speaking, stars between Rin and Rout are most probably cluster members although some of these stars might still be field stars and some of the stars that are not between Rin and Rout might still be cluster members. This was the first technique we adopted to get an improved CMD with fewer field stars and with fewer stars that have big photometric errors. The locations of Rin and Rout on the cluster’s image are show in in Figure 5.3.

At this point, the best representation of the cluster’s CMD is part (c) of Figure 5.4. However, more correction for the field stars is needed. To do so, we took 10 field images that are around 5 arcminutes away from the cluster center. Four of these images were in the V filter (80s exposure), 3 in the IL and 3 in the IS. We processed, combined and did calibrated photometry on these images just as we did with the cluster images. We did this to plot a CMD of this part of the sky. We then selected an area from the field images that is equal to the area of the best region in our cluster ($Rin < R < Rout$). Finally, we used a program written by Dr. Andrew Layden that removed stars from our improved CMD with similar magnitude and color to stars in the field CMD (see Figure 5.4). The program looks
Figure 5.2: CMD of NGC 6494. Part a of this diagram includes all of the stars we detected on the CCD. Part b represents stars that are close to the center of the cluster. Part c represents all of our stars detected on the CCD camera minus stars close to the center and stars away from the center. Part d represents stars that are far from the center.

at \((V - I)\) color and at the \(V\) magnitude of each star in the CMD of the field stars. Then it tries to find stars with similar magnitudes and colors on the cluster’s CMD and subtracts them from the cluster’s CMD. In general, the field star’s \((V - I)\) colors ranged between 0.5 and 2 and the \(V\) magnitude ranged between 12 and 17.5. These were reasonable values for the field stars especially because we are not concerned about stars with faint magnitudes because in this study we are more concerned about stars with bright magnitudes (LPVs), which is why the small telescope we used was sufficient for this study. Statistically speaking, our final CMD diagram after the subtraction contains fewer field stars.
Figure 5.3: Rin (red) and Rout (green) plotted on our cluster’s image. All stars between Rin and Rout are probably members of the cluster.

5.1 Interstellar Reddening

NGC 6496 is 4.2 kpc away from the galactic center; the galactic longitude and latitude in degrees for our cluster are 348.03 and -10.01, respectively [24]. Consequently, light coming from the cluster must be passing through gas and dust (interstellar medium) before we detect it, which causes it to become redder and fainter. This reddening occurs due to the light scattering off dust. Usually the interstellar medium scatters light with shorter wavelengths (blue) more than it absorbs longer wavelengths (red), that is where the “reddening” name comes from. In other words, the colors of the stars we detected are affected by the interstellar medium and thus we call them the observed colors. We refer to the real colors of the stars as the intrinsic colors, which are the colors of the stars if reddening didn’t occur. The relation
Figure 5.4: (a) CMD of all of the stars detected on our CCD for $12.5 \leq V \leq 17.5$. (b) CMD of stars detected in our field image. (c) Our improved CMD (d) Our best and final CMD ((c)-(b))

between the observed and the intrinsic colors is not complicated and is expressed in equation 5.1:

$$E(V - I) = (V - I)_{obs} - (V - I)_{int}. \quad (5.1)$$

where $E(V - I)$ represents the reddening, $(V - I)_{obs}$ represents the observed color and $(V - I)_{int}$ represents the intrinsic color. The reddening in the (B-V) color was found in the literature to be $E(B - V) = 0.15$ [24]. We used the following formula $E_{V-I} = E_{B-V} \times 1.24$ that was provided by Layden *et al.* [25]. Hence our reddening was found to be $E(V - I) = \ldots$
We can apply this shift to the color of the cluster, as shown in equation 5.2:

\[(V - I)_{obs} = (V - I)_{int} + E(V - I).\]  \hspace{1cm} (5.2)

Now that we calculated the reddening value, we still need to shift the V magnitude of our stars using the apparent visual distance modulus \((\mu_{app})\) expressed in equation 5.3:

\[\mu_{app} = V_{obs} - M_v\]  \hspace{1cm} (5.3)

where \(M_v\) is the absolute magnitude of the star and \(V_{obs}\) is the observed magnitude of the star including extinction. In order to relate the observed magnitudes to absolute magnitudes for our cluster, we used Harris’s value of the apparent visual distance modulus, \(\mu_{app} = 15.74\). Thus,

\[V_{obs} = \mu_{app} + M_v.\]  \hspace{1cm} (5.4)

### 5.2 Isochrones

In order to test the reddening value from Harris, we chose to use isochrones to see whether they fit our CMD after applying the shifts from the previous section. We used isochrones provided by Girardi et al. [26] for an age of \(T=11.2\) Gyr. We tried isochrones with different metallicities ranging from -2.3 to 0.2 dex. The two isochrones that best matched our CMD are shown in Figure 5.5. The two isochrones shown have metallicities of -0.4 and -0.7 dex, bracketing our cluster’s metalicity of -0.46 dex [24].

We shifted the isochrones horizontally by 0.186 (for reddening) and vertically by 15.74. Looking at Figure 5.5, one can see that the blue isochrone with metallicity of -0.4 dex fits our CMD better than the green isochrone of metallicity -0.7 dex. One can see that our cluster’s RGB falls on the RGB of the blue isochrone (dashed line) while it’s redder than the
RGB of the green isochrone. Also, the AGB of the isochrone with -0.4 dex metallicity (blue solid line) agrees better with our CMD’s AGB than the AGB of the isochrone with -0.7 dex metallicity (green solid line). This is a confirmation that our cluster’s metallicity is close to -0.4 and is more metal rich, which means that it is redder than a -0.7 dex. 5.5.
Figure 5.5: Two isochrones models on our CMD. The green lines represent the model with metalicity of -0.7 dex. The blue lines represent the model with metalicity of -0.4 dex. The isohchrone RGBs are shown as dashed lines, and the AGBs as solid lines.
CHAPTER 6
VARIABLE STARS

Variable stars were detected from the two different photometric processes used (ISIS and DAOPHOT/ALLFRAME). The criteria used in identifying variable stars in ISIS are quite different from the criteria used to identify variable stars with DAOPHOT and so detecting variable stars with both methods makes us more confident about identifying these stars.

6.1 Variable Stars with ISIS

As discussed in the chapter 3, ISIS uses a subtraction method which leaves us with an image that contains signals from variable stars (see Figure 6.1). We examined all of the bright signals that appeared with ISIS in this image. The locations of the signals on the cluster image are shown in Figure 6.2.

Looking at Figure 6.1 and seeing a lot of “white” signals, one might think there should be more variables stars (green circles) on Figure 6.1. However, it was hard to distinguish between real variable stars and noise. Because of that, we examined all the signal found on the subtracted image and we selected the variable stars according to a threshold and according to how well the signal is fitted. We observed all the signals on the var.fits image and we looked at their radial profiles to see how well they are fitted. Most of the signals detected were poorly fitted and had a lot of scatter. Moreover, we tried plotting the light curves of the stars associated with the signals to see if there was significant variation in their fluxes or if their fluxes were constant. Based on this, we were able to separate variable stars from non-variables using ISIS. In Figures 6.1 and 6.2, there is a star that was marked with both a green and a red circle almost in the center of the image. This star was not considered to be a variable star although it had a large signal. This is a bright star that saturated the CCD. After examining these figures, we identified 5 variable stars with ISIS (marked with green circles in Figure 6.2).
Figure 6.1: Subtracted image from ISIS. All the signals left after applying the subtraction method are marked with circles. Green circles represent possible variables stars. Stars marked with red circles have a signal that falls below our threshold intensity and/or they have bad radial profile fitting.

6.2 Variable Stars with DAOPHOT

We used the variability with the DAOPHOT photometry to distinguish between a variable and a regular star. Unlike ISIS, DAOPHOT provides us with magnitudes for each star and so we were able to look at the variation of the magnitude in order to tell if the star was variable or not. We define the variability index (∆) to be the ratio of the standard deviation (σ) of the star’s magnitudes and the error generated from DAOPHOT:

$$\Delta = \frac{\sigma}{\text{error}}$$  \hspace{1cm} (6.1)

The standard deviation represents how much the magnitude of the star varies from the
mean magnitude in each filter. If the star is varying in magnitude over the time, then the standard deviation should be large and so should the variability index in the condition that we have small errors. If the star is non-variable, then the standard deviation should be small and the variability index will be small, too. Plotting the magnitude of the stars versus the variability index can help us identify possible variable stars in our cluster (see Figure 6.3).

All of the stars with variability indices greater than 4 (around 30 stars) are possible variables and hence were investigated. Having a large variability index doesn’t directly mean that the star is variable because a large variability index can occur for bright stars, blended stars in crowded regions, binary stars, etc. Plots of the variability index can also give us a sense about the type of the variable star by examining the magnitude/color of the star and the value of the variability index. For example, a red star with a large variability index is most probably a Long-Period Variable Star (LPV) while a blue star with a lower
variability index is most probably a short period variable star. We plotted the light curves\textsuperscript{1} of all the suspected variables and we looked at their radial profiles in order to distinguish between real variables and non-variables. Most of these 30 suspected variables were actually non-variables. Many of them appeared as variables because they were blended with other stars on nights with bad seeing. When two stars are detected as one star, its magnitude will be brighter. When these stars are detected on another night as two separate stars, their magnitudes will be fainter. This causes a variation in magnitude, which could lead to a large variability index. At this point, we have our variable star candidates and we plotted the light curves to identify the true variable stars, find their periods and classify them.

6.3 Period Detection

More than one method was used to detect the periods of our variable stars. Because most of the long-period variable stars that we found are irregular and because we don’t have

\textsuperscript{1}magnitude vs time
many cycles, it was hard to determine a definite period. However, detecting periods with
different software and comparing them to see if they agree give us a sense of how reliable
our periods are.

### 6.3.1 Phase Dispersion Minimization (PDM)

The first method used to determine the periods for long and short period variables was
PDM. PDM divides the light curve into phase bins and tries different periods and calculates
the scatter of the data about the mean light curve. PDM uses a statistical quantity (theta, θ),
the ratio of the weighted scatter in the bins to that for the whole light curve, (to evaluate
the periods) [34]. Having a small theta (near zero) means that the variation is regular, the
weighted sum of the scatter in the bins is much less than the scatter for the whole light
curve, and thus this is the true period. On the other hand, having a theta close to 1 means
that that variation is irregular [28] and that the weighted sum of the scatter in the bins is
comparable for that in the whole light curve [34]. PDM tries different periods and calculates
theta for each period, shown on a figure (see Figure 6.4). Most of the time, we chose the
period at the minimum because it represents the period with the best fit.

### 6.3.2 Method of Template Fitting (MTF) by Dr. Andrew Layden [31]

Another program written by Dr. Andrew Layden [23] was adopted to detect periods of
variable stars. This program uses ten carefully chosen templates of light curves. Different
templates represent light curves for different types of variable stars (RRab, RRc, LPV,
binaries, etc.). This method is powerful because it handles different types of variable stars.
The data are folded by a set of periods ranging over the expected values and χ² values are
calculated [32]. As shown in equation 6.2, χ² represents the deviation of each point from the
template and so the minimum χ² represents the best fitted template.
Figure 6.4: Variation of theta with the period. The period that corresponds to the minimum theta is the most probable period of our light curves.

\[ \chi^2 = \frac{\sum_{i=1}^{N}(Y_i - Y_{fit})^2}{(N - 1)} \] (6.2)

The program takes the input templates, which have an amplitude of 1.0, and it does a least squares fit to the data that scales the template to fit in amplitude, shifts the template vertically to give a mean mag, and shifts it horizontally to give a phase shift and computes \( \chi^2 \) and rms around the best fit. After we choose our period from Figure 6.5 (which most of the time corresponds to the minimum \( \chi^2 \)), the program plots the phased light curves at our chosen period on the best four fitted templates (see Figure 6.6). Then, we choose the best fitted template out of the best four based on the \( \text{rms} \) error and we end up with a value for the period, a magnitude vs. phase graph (see Figure 6.7), the rms error, the range of
magnitudes, the best-fit amplitude, the intensity-mean magnitude and the S/N ratio (which we call PI). These results for our 6 LPVs are given in table 6.1.

**Range $\Delta V$, Amplitude $A_V$ and S/N (PI) in MTF**

Range and amplitude are very similar for a light curve of regular behavior. However, these two terms can be very different for irregular or semi-regular light curves. The range $\Delta V$ or $\Delta I$ represents the difference between the maximum and the minimum magnitude reached during our whole observing time in the V and I filters, respectively. If the pulsation was regular, then the light curve should reach the same minimum and maximum in each cycle. On the other hand, $A_V$ is the fitted amplitude of the template curve. So if the pulsation of the star is very regular (i.e. Mira stars), the amplitude and the range should be
Figure 6.6: Phased light curves at our chosen period for the four best fitted templates.

similar because the minimum and the maximum of the light curve and the template should be close as the pulsation would be regular. If the star was completely irregular, we expect the amplitude and the range to be very different as the minima and maxima of the of the irregular light curve will change from cycle to another while they will remain the same in the template. So, comparing the range and the amplitude will help us classifying the types of our variable stars.

Finally, the S/N (PI) represents the fitted amplitude of the template divided by the rms of the points around the template. Having a large amplitude with a small rms means that the template is well fitted and that the period is regular i.e. the star is more probably a Mira star. On the other hand, if the template didn’t fit the light curve pretty well, then we will
Figure 6.7: Our final Phase Vs. Mag plot of one of our variable stars. Notice how the points are well fitted to the template for the given period of 71 days.

have a large rms and consequently a larger PI. This can be an indication that the variation of the light curve is irregular.

6.4 Long-Period Variable Stars

We detected 6 LPVs and estimated their periods. All of these stars have a red color which make us more confident about our classification. They also have an irregular period ranging from 60 to 350 days. In the following subsections, we discuss the properties of each LPV and we provide their light curves.
Star ID-31

Star ID-31 is 2.08 arcminutes away from the center of the cluster. It is red in color ($< V > - < I > = 3.21$) and it belongs to the red giant branch on the CMD (see Figure 6.22). It was detected as a variable star with ISIS and DAOPHOT (using the variability index with DAOPHOT and the signal from applying the subtraction method with ISIS). This star is relatively bright with a magnitude varying between 14.4 and 15.5 in the V filter. Also, one can see more than one of the star’s semi-regular cycles (see Figure 6.8$^2$). One can see some periodicity in the variation. Because the periodicity is not quite regular, we can classify this star to be a semi-regular LPV.

Figure 6.8: Light Curve of one of our LPVs. Better than average seeing magnitudes are represented by circles and worse than average seeing magnitudes are represented by crosses. We gave more weight for magnitudes marked with circles and less weight for magnitudes marked with crosses when the period was estimated.

$^2$Time is expressed in partial Julian date. pJD. pJD = JD-2450000.
In Figure 6.8 one can see that the variations of magnitude in all three filters agree with each other because the magnitudes reach maximum and minimum at the same time in the IL, IS and V exposures, which makes us more confident that this variation in magnitude is due to the star’s pulsation rather than instrumental or observational effects. We see the same result in the flux vs. time plot (see Figure 6.9), where the fluxes represent the number of counts in the individual ISIS difference images.

Finally, we measured the period of this star using PDM and MTF and the data from DAOPHOT and ISIS. Using PDM on the light curves from DAOPHOT in the V, IL and IS observations we found the period to be 70, 71.5 and 70 days, respectively. Using PDM on the data from ISIS in V, IL and IS we found the period to be 70.4, 70 and 70 days, respectively. MTF was used to find the period of the light curves generated from DAOPHOT and the
period was found to be 71, 70 and 70 days in V, IL and IS, respectively. The mode value for all of the calculated periods was 70 ±0.55 days. The uncertainty of the period represents the standard deviation divided by the square root of the number of periods, 9 in this case. This is the technique we used to calculate the periods for all of our variable stars. This period seems reasonable based on visual inspection of the light curve. The phased light curve in the V filter of this star using MTF (with a period of 71 days) is shown in Figure 6.7. The scatter is relatively small and the data points fit the model pretty well.

The range and the amplitude of this star are 1.24 and 0.50, respectively, in the V filter. The range of 1.24 tells us that the magnitude of the star was really changing and that the change was relatively large, a characteristic of LPVs. The amplitude being significantly smaller than the range indicates that the maxima and minima of the light curve were changing with every cycle, which makes us believe that this is not a Mira star. If the maxima and minima of the light curve were not changing in each cycle, it would have meant that the variation was regular and that the amplitude was close to the range, indicating a possible Mira star.

Finally, the PI value of 3.13 can tell us how the fitting of the light curve went. A large value of PI means that the fitting of the light curve went well, an indication of a regular light curve (i.e. a MIRA). A small value of PI means that the fitting of the light curve was poor, an indication of an irregular light curve (i.e. IRR, SR, etc.). However, there is no standard value for a large or a small PI since every light curve can behave differently depending on the type of the variable star. All of the analysis of this star led us to believe that this is a semiregular LPV.

**Star ID-32**

Star ID-32 is 2.21 arminutes away from the center of the cluster. It is red in color (<V> – <I> = 3.12) and it belongs to the red giant branch on the CMD (see Figure
It was detected as a variable star with ISIS and DAOPHOT. Figure 6.10 shows that the light variations in all three observations are consistent (reach same minima and maxima at the same time). The magnitude varies between 14.6 and 15.3 in V.

Figure 6.10: The light curve of star ID-32. This star is a semiregular LPV as the variation of the magnitude in IL, IS and V is not following a definite pattern, although it is varying. Better than average seeing magnitudes are represented by circles and worse than average seeing magnitudes are represented by crosses.

In addition to DAOPHOT, ISIS classified this star to be variable as well (see Figure 6.11). This star’s range and amplitude are 0.74 and 0.44 in the V filter, respectively, and its PI was found to be 3.002. The range shows a real variation of the star’s magnitude (i.e. it is a variable star) and the amplitude is relatively close to the range, but not the same, which means that the light curve has similar minima and maxima in each cycle and this shows
some regularity. Also, the PI value tells us that the scatter was small and that the fitting went well, producing a small rms and a relatively large PI.

The phased light curve of the star is shown in Figure 6.12. The data points are well fitted by the template. Looking at this fit, at the shape of the light curve and at the values of the range, amplitude and PI we concluded that this is a semiregular LPV with a period of $107 \pm 1.2$ days. All of the calculated periods using the different period-finding methods in all the filters gave us periods between 105 and 109 days.

![Plot 1 = ID-32](image)

Figure 6.11: The flux variation of star ID-32 over time. One can see the semiregular periodicity, which agrees well with the light curves from DAOPHOT.

**Star ID-29**

ID-29 is another LPV found in our cluster 2.64 arcmin away from the center and it is also on the red giant branch of the CMD (see Figure 6.22). It has a red color ($< V >$ - <
Figure 6.12: The phased light curve of ID-32, relatively well fit by the template for a semi-regular.

$I \geq 2.4)$ and the magnitude varies between 13.9 and 14.7 in V. The range, amplitude and PI in the V filter are 0.64, 0.42 and 3.603, respectively. The range and the amplitude of the star are relatively close, which tells us that there is some regularity in the light curve. The period of this star using PDM and MTF on the light curves was $350 \pm 8.2$ days. However, by looking at the light curve (see Figure 6.13) and Figure 6.14) we concluded that the period of 350 days is a long secondary period. Because irregular LPVs pulsate in more than one mode, they usually have one secondary period (longest) and other shorter periods. In order to investigate this further and to confirm that this is the long secondary period, we would need to observe this cluster over a longer period of time. If the possible long secondary period is 350 days then we would need observations over more than a thousand days. We decided to manually find the shorter pulsation period of this star. We calculated the time between each two consecutive peaks and the average of these times to represent the shorter period of this star. The standard error is the standard deviation divided by the square root of the number of cycles that we counted. The result was $43 \pm 4.4$ days.
In summary, we estimate the long secondary period of this star to be 350 ± 8.2 and its short pulsation period to be 43 ± 4.4 days. The range, amplitude, PI and the light curve indicate that this probably a semiregular LPV.

![Star ID-29 light curve](image)

Figure 6.13: The light curve of star ID-29. Better than average seeing magnitudes are represented by circles and worse than average seeing magnitudes are represented by crosses.

**Star ID-30**

The LPV star ID-30 has also a red color ($<V> - <I> = 3.37$) and belongs to the red giant branch on the CMD (see Figure 6.22). It was detected with both of our photometric processes and it is relatively bright. Its V magnitude was oscillating between 14.7 and 15.5. It is not a crowded star and is 1.14 arcminutes away from the cluster center. From the light curve of Figure 6.15, we can classify this star as an irregular LPV. There is definitely a variation in magnitude with some periodicity that is irregular.
Figure 6.14: The flux variation of star ID-29 over time. One can see the periodicity of the magnitude, which agrees well with the light curve from DAOPHOT.

Magnitudes of star ID-30 in the V filter were not detected in the first season, probably due to photometric error from over exposure. However, the variations of the magnitudes in all three filters are consistent, reaching minima and maxima at the same time. In addition, the flux vs. time variation is shown in Figure 6.16. The range in the V filter was found to be 0.62 and the amplitude is 0.19 in the same filter. The PI value is 1.34, which is 1.79 less than the semiregular LPV ID-31. The range in this case gives us confidence that there is some variation in the magnitude of the star. The values of amplitude and PI indicate that the rms was large and that the fitting wasn’t as good as for star ID-31. This must have happened because the light curve for this case is irregular. For this irregular star, different period-finding methods gave us different periods. For instance, using MTF the period was about 90 days while it was estimated to be 73 days when using the PDM method. The
Figure 6.15: The light curve of star ID-30. One can see the irregular behavior of the variation of the magnitude in IL, IS and V over a period of almost a 100 days. This makes us classify this star to be an Irregular LPV. Better than average seeing magnitudes are represented by circles and worse than average seeing magnitudes are represented by crosses.

Phased light curve for the period of 91.5 days is shown in Figure 6.17. One can see that the data points don’t fit the template very well, which make us question if the period of 91.5 days is correct. PDM gave us a period of 73 days so we plotted the phased light curve with a period of 73 days (see Figure 6.17). The fit was improved with the period of 73 days. For further confirmation, we ran PDM on our data from ISIS and we found a period of 75 days. We chose to adopt the period that came from applying PDM, 73±6.4 days. It is nearly impossible to be certain of the period because its irregular and the period is long. The irregularity of this star is indicated not only by the appearance of the light curve, but
also by the disagreement of the periods we found and the appearance of the phased light curve. Comparing part (b) of Figure 6.17 and Figure 6.7 one can see that the light curve for the semiregular ID-31 star is smoother and better fitted than that of the plot of the irregular ID-30.

![Plot6 = ID-30](image)

**Figure 6.16:** The flux variation of star ID-30 with time in all of our three observations.

**Star ID-27**

We found that ID-27 is also an LPV. Just like the other LPVs, this star is on the red giant branch of the CMD ((see Figure 6.22) and it is red in color ($< V > - < I > = 2.82$). The magnitude of this star varies between 14.2 and 15.0 in V. The star’s range in the V filter was 0.64 and its amplitude in the same filter was 0.12. The value of the range tells us that the magnitude of the star was changing with time, but the amplitude much smaller, which indicates the irregularity of the light curve. Its PI value of 1.198 indicates that the
template wasn’t fitted as well as our other semiregular LPVs due to the irregularity of its light curve, as seen in Figure 6.18. Thus, we believe that this is an irregular LPV. The period was found to be $94 \pm 4.2$ days. Note that this star was detected as variable in ISIS as well as in DAOPHOT. The variation of the flux for this star is plotted in Figure 6.19. Since this is an irregular LPV, we cannot be certain about the exact value of the period, but we believe that 94 days is the most probable period based on our observation.

**Star ID-18**

Another LPV found was ID-18. This red star ($<V> - <I> = 2.33$) belongs to the red giant branch of the CMD (see Figure 6.22). Although the behavior of this star is consistent in IL, IS and V, ISIS didn’t indicate this as a variable star. This could be because star ID-18 is further from the center of the cluster (3.85 arcminutes) and as discussed in chapter 3, ISIS has an advantage when working near the center but a disadvantage when working away from the center. Although the PI of this star is relatively high, 3.628, the range and amplitude in the V filter were relatively low with values of 0.39 and 0.32, respectively.
Figure 6.18: The light curve of star ID-27. This star has a period of 94 days. All of the light curves in our three observations agree well. Better than average seeing magnitudes are represented by circles and worse than average seeing magnitudes are represented by crosses. The PI value indicates the light curve is relatively well fitted, while the small amplitude and range indicates this is an irregular LPV. Looking at the light curve of this star in Figure 6.20 and at the phase plot in Figure 6.21, and considering the range and amplitude, we classified this star as an irregular LPV with a period of 47±0.9 days.
Figure 6.19: The variation of the flux over time for star ID-27.
Figure 6.20: The light curve of star ID-18. This star has a period of 47 days. All of the light curves in our three observations agree well. Better than average seeing magnitudes are represented by circles and worse than average seeing magnitudes are represented by crosses.
Figure 6.21: The phased light curve of star ID-18.
Figure 6.22: CMD of NGC 6496. LPVs and their ID numbers are marked with pink stars.
Table 6.1: LPVs

X and Y are the coordinates on Figure 6.2. R is the distance from the cluster. $<V>$ and $<I>$ are the mean V and I magnitudes. $R_v$ and $R_i$ are the range in the V and I filters, respectively. $A_v$ and $A_i$ are the amplitude in the V and I filters, respectively. $N_{cycle}$ represents the estimated number of cycles that we observed for each LPV. $N_{vobs}$ and $N_{iobs}$ are the number of frames we had in the V and I filter respectively. S/N represents the fitted amplitude of the template divided by the rms.

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<th>ID</th>
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<th>Y</th>
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<th>$&lt;I&gt;$</th>
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<th>$R_i$($\Delta I$)</th>
<th>$A_v$</th>
<th>$A_i$</th>
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<td>47</td>
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$^a$Semi-Regular LPV

$^b$About 9 cycles for the short period of 43 days and 2 cycles for the long secondary period of 350 days

$^c$Long secondary period: 350 days

$^d$Irregular LPV
CHAPTER 7
SHORT-PERIOD VARIABLE STARS

In order to find periods of short-period variable stars, we would like to have many observations separated by equal and short intervals of time (e.g., minutes, hours, or days). In our case, we were imaging the cluster every two weeks. So it was easy to detect variation in brightness of stars (including short-period variable stars) but it was difficult to find the period of short-period variable stars because of the relatively long time between observations. Although this research focuses on LPVs, we were able to find 5 short-period variable stars which are all eclipsing binary stars. We used all the data we have for these short period variable stars plus data for this cluster that was acquired by Dr. Andrew Layden in 1996 [23]. Having two different data sets from two different telescopes and using different period-finding techniques helped us in detecting the period and types of these stars. Due to the long time between observations, we give our best estimate of the periods and classification for each star based on their light curves, periods, amplitude etc.

7.1 Star ID-256

Star ID-256 is 4.21 arcmin away from the cluster center. It has a range of 0.30 and an amplitude of 0.23 in the V filter. The range indicates a change in the magnitude of the star, which can also be seen by looking at the light curve in Figure 7.1. The amplitude and the range are close, which indicates a good periodicity in the light curve, that the light curve was reaching the same minima and maxima in every cycle.

To find the period of this star, we used PDM and tried a wide range of periods, and we selected different periods that corresponded to the smallest values of theta. In PDM, we calculated the scatter for these different periods and found that the best fit (minimum scatter) corresponded to a period of 0.76664 day. We used MTF for the best periods found with PDM and again found that the best fit was for a period of 0.76664 day. The best fitted
Figure 7.1: The light curve of star ID-256. Better than average seeing magnitudes are represented by circles and worse than average seeing magnitudes are represented by crosses.

The template for this star was the W UMa template used in MTF. The phased light curve in Figure 7.2 shows two nearly equal minima in each cycle, which indicates that this star is a W UMa star. The minima are almost the same in each cycle, which is a characteristic of W UMa stars as explained in chapter 1. The two stars have nearly the same surface temperature and size. So we see nearly the same minimum brightness when either of them is eclipsing the other. The change in magnitude in the V, IS and IL exposures is the same, about 0.3, which is an indication that these stars are not pulsating but are eclipsing each other.

Finally, we fitted the light curve to different periods that are slightly more and less than 0.76664 days and we found that the best corresponds to a period of 0.76664± 0.00005 day. Considering the similar values of the range and the amplitude, the best-fitted periods given
Figure 7.2: The phased light curve of star ID-256. One can see two near equal minima in each cycle, corresponding to the two different eclipses these stars undergo in every cycle. This best-fitted template corresponds to a W UMa star. One can see the small scatter of the folded data, which indicates that the period is relatively accurate. The phased light curve for the V exposure is represented with the green color. The phased light curve for the IL and IS exposures are represented with the black and red color, respectively.

by PDM and MTF, the phased light curve, the wide range of periods that we tried on this star and at the best fitted template in MTF, we conclude that this star is an W UMa star with a period of $0.76664 \pm 0.00005$ day.

7.2 Star ID-244

Star ID-244 is 5.89 arcmin away from the cluster center, relatively far. The light curve of this star in Figure 7.3 shows constant drops of magnitudes, which can be a result of some eclipse phenomenon.

The range and the amplitude in the V filter were 0.48 and 0.40, respectively. The
Figure 7.3: The light curve of star ID-244. One can see the constant drops in magnitudes in all three exposures, a behavior of eclipsing binary stars. Better than average seeing magnitudes are represented by circles and worse than average seeing magnitudes are represented by crosses.

similarity of the range and the amplitude indicates that the minima and maxima were almost the same in every cycle. Just like we did with star ID-256, we used PDM and investigated a wide range of periods until we found a few possible periods for which the phased light curve was well fitted. Then, we used MTF with the best ranges of periods we found with PDM. The best fitting template was a W UMa template at a period of 0.28974 day. To investigate this further, we plotted the phased light curve with our best estimated period (see Figure 7.4). The light curve shows a small scatter. One can also see that the values of the minima in each cycle are very close to each other, a characteristic of W UMa stars.

The similarity of the range and the amplitude, the best fitted template with MTF, the two nearly equal minima in each cycle, the small magnitude variation and the shape of the
Figure 7.4: The phased light curve of star ID-244. One can see two nearly equal minima in each cycle, corresponding to the two different eclipses these stars undergo in every cycle. This best-fitted template corresponds to a W UMa star. One can see the small scatter of the folded data, which indicates that the period is relatively accurate. The phased light curve for the V exposure is represented with the green color. The phased light curve for the IL and IS exposures are represented with the black and red color, respectively.

phased light curve are all indications that this is a W UMa variable star with a period of 0.28974± 0.00001 day.

7.3 Star ID-115

Star ID-115 is 4.95 arcmin away from the center of the cluster. The range and the amplitude of this star are 0.42 and 0.59, respectively, on the IL exposures. The light curves in Figure 7.5 show that there is variation of magnitude with time in all three exposures. However, because this star is near the edge of the CCD chip, we were not able to observe it on most of our nights. Hence, it was difficult to detect an accurate period for this star.
Since we have small number of observations, we used Dr. Layden’s observations of this cluster with the 0.9m telescope in 1996. He found a period of 0.3365 day. Just like we did with the other short-period variable stars, we tried various periods, (not only around 0.3365 day), and found the best fit at a period of 0.3371 day. This period was found with MTF by fitting the W UMa template. Although we have few data points, the phased light curve in Figure 7.6 shows a good periodicity and a small scatter.

The range and the amplitude of this star are relatively close. The change in magnitude in the V, IS and IL exposures is the same, about 0.4, an indication that this star is not pulsating but eclipsing. Moreover, in each cycle there are two near equal minima. Finally, the minimum $\chi^2$ value was found by fitting the light curve to a W UMa template. All of these
Figure 7.6: The phased light curve of star ID-115. Although we have small amount of data for this star because of its position on our CCD, one can still see two near equal minima in each cycle. The phased light curve for the V exposure is represented with the green color. The phased light curve for the IL and IS exposures are represented with the black and red color, respectively.

things lead us to conclude that this is a W UMa star with a period of $0.3371 \pm 0.0002$ day. For this star, more observations are needed, preferably with short intervals for a better and more accurate period to be determined.

7.4 Star ID-418

Star ID-418 is 2.60 arcmin away from the cluster center. The range and the amplitude are 0.33 and 0.30, respectively in the IL exposure. The similar values of the range and the amplitude is an indication that the star was reaching the same minima and maxima in each cycle. Just like star ID-115, we don’t have many data for this star (see Figure 7.7) so it was
difficult to determine a precise period. Just like we did with the other short-period variable stars, we tried a wide range of periods. Using PDM, we found the best fit at a period of 0.884 day. We then took this period and tried it with MTF, which gave the best fit with a W UMa template, a good indication that this star is of that type.

Figure 7.7: The light curve of star ID-418. One can see changes in magnitude over the time in all three exposures. We were not able to observe this star in most of our nights. Better than average seeing magnitudes are represented by circles and worse than average seeing magnitudes are represented by crosses.

Moreover, one can see two nearly equal minima in each cycle. The phased light curve in Figure 7.8 shows a small scatter, indicating a relatively accurate period. The V photometry for this star was incomplete due to photometric errors, so it is not shown in the phased light curve. The range and the amplitude of this star were relatively close. Moreover, in each cycle we have two near equal minima. Finally, the minimum $\chi^2$ value was found by fitting the light curve to a W UMa’s template with MTF. So we believe this is a W UMa star with
a period of $0.884 \pm 0.001$ day. For this star, more observations are needed, preferably with short intervals for a more accurate period to be determined.

![Graph](image.png)

Figure 7.8: The phased light curve of star ID-418. Although we have small amount of data for this star because of its position on our CCD, one can still see two near equal minima in each cycle. The phased light curve for the IL and IS exposures are represented with the black and red color, respectively.

### 7.5 Star ID-226

Star ID-226 is 4.81 arcmin away from the cluster center. The range of this star is 0.87 in the I filter and the amplitude is 0.59 in the same filter. The light curve seen in Figure 7.9 shows the same magnitude drop in all filters, an indication that this could be a binary star and these drops correspond to the time when the stars eclipse each other.

To find the period of this star, we used PDM with a wide range of periods and at a period of 0.43851 day. We used MTF on different ranges of our best periods found in PDM.
and again found the best fit at a period of 0.43851 day. The best fitted template in MTF was the Algol template. The two different magnitudes for the minima in each cycle also indicate this star could be an Algol variable. Unlike W Uma stars, these stars are usually of different sizes, temperatures and masses, which is why we see two minima with different magnitudes in each cycle.

Considering the similar values of the range and the amplitude, the periods given by PDM and MTF and the small scatter and two different minima in the phased light curve, we believe that this star is an Algol variable star.
Figure 7.10: The phased light curve of star ID-226. One can see two different magnitudes for the minima in each cycle. The light curve for the V exposure is represented with the green color. The phased light curve for the IL and IS exposures are represented with the black and red color, respectively.

<table>
<thead>
<tr>
<th>ID</th>
<th>X</th>
<th>Y</th>
<th>R(arcmin)</th>
<th>Period (days)</th>
<th>error (days)</th>
<th>Type</th>
</tr>
</thead>
<tbody>
<tr>
<td>256</td>
<td>930.60</td>
<td>334.67</td>
<td>4.21</td>
<td>0.76664</td>
<td>0.00005</td>
<td>W UMA</td>
</tr>
<tr>
<td>244</td>
<td>61.91</td>
<td>841.29</td>
<td>5.89</td>
<td>0.28974</td>
<td>0.00001</td>
<td>W UMA</td>
</tr>
<tr>
<td>115</td>
<td>1009.24</td>
<td>543.81</td>
<td>4.95</td>
<td>0.3371</td>
<td>0.002</td>
<td>W UMA</td>
</tr>
<tr>
<td>418</td>
<td>358.14</td>
<td>168.16</td>
<td>2.60</td>
<td>0.884</td>
<td>0.001</td>
<td>W UMA</td>
</tr>
<tr>
<td>226</td>
<td>840.96</td>
<td>808.75</td>
<td>4.81</td>
<td>0.3371</td>
<td>0.0002</td>
<td>Algol</td>
</tr>
</tbody>
</table>

Table 7.1: Short-Period Variable Stars Table.

X and Y are the coordinates, R is the distance from the cluster center.
Figure 7.11: CMD of NGC 6496. LPVs and their ID numbers are marked with pink stars. Short-period variables and their ID numbers are marked with blue filled squares.
CHAPTER 8
CONCLUSION

In this study, we searched for variable stars in the metal-rich globular cluster NGC 6496. We imaged the cluster with two different filters once every two weeks for a total time of about two years. We used three types of calibration images (bias frames, dark frames and flat fields) to help reduce the noise and the defects in the CCD and the filters for a more reliable data. We used two different programs to do our photometry (ISIS and DAOPHOT). ISIS works on a subtraction method while DAOPHOT uses the aperture photometry method.

Using two different methods to do our photometry gave us a chance to find variable stars in two different ways using two different criteria. To calibrate our magnitudes, we compared three different sets of standard stars and we ended up using the best set (by Stetson) to calibrate our stars.

In total, 11 variable stars were found. Six of those variable stars are Long Period Variable Stars and are new discoveries and the rest (5 Stars) are short period variable stars in which 4 of them are W UMa binary stars and one is an Algo binary star. We were able to detect the periods of these variable stars and classify them according to their light curves, phased light curves, periods, color, amplitude and range. We also compared our light curves with other lightcurves’ templates that correspond to different types of variable stars using MTF software. We plotted the CMD of our cluster and located the locations of our variable stars, which also helped us in classifying the type of the variable stars (especially LPVs). The main reason for why we are interested in studying the properties of LPVs is that they are at their last phase of evolution.

Since this study focuses on LPVs and these stars can have pulsation periods that are more than a thousand day and since we observed these LPVs for around two years, it would be very helpful to observe this cluster for a longer amount of time. This will help us detecting more accurate periods for the stars and possibly be able to find the different periods for the
REFERENCES


[8] This AAVSO Variable Star of the Season article was written by Kerri Malatesta.


[16] From Dr. Andrew Layden


different pulsating modes in these stars. In addition to that, a Period-Luminosity relationship can be plotted for the LPVs to try to understand how LPVs climb on the RGB. Finally, a more accurate study can be done on studying the membership of our variable stars to the cluster.


[23] From Dr. Andrew Layden.


[33] Glossary of Meteorology: Radiative Flux

[34] University of Nebraska-Lincoln astronomy education http://astro.unl.edu/naap/vsp/pdm.html