

# Stellar CCD Photometry

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## Abstract

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Our Final goal in this experiment is to create a color-magnitude diagram of the open cluster M52. Using this diagram and theoretical models of stellar evolution we will determine the distance and the age of the cluster.

## 1 Telescope

We will use a 70 *cm* telescope located at the MPIA. The CCD detector in the telescope is a Loral/Lesser n2k2eb BI thinned, back illuminated CCD with 2048 x 2048 pixels with a size of 15  $\mu m$ . For our measurement we use a binning factor of two which means four pixels (two in each direction) will be regarded as one speeding up the read out process.

A CCD (Charge Coupled Device) is a detector that detects photons. The photoelectric effect produces a charge that is stored for every pixel. These charges integrated over the exposure time are then read out line by line producing an image through an analog digital converter. We use the IRAF toolkit for our analysis.

Since we need measurements of magnitudes of stars in two different bands, we use the Johnson Filters B and V.

There are multiple reasons to use a CCD detector for stellar observations.

- Quantum efficiency: The CCD has got a very good quantum efficiency of up to 90% (depending on the wavelength), that is important since we do not receive much light from faint sources commonly.
- Linearity: The signal we measure is linear dependent on the number of photons hitting our detector, so we can easily determine the brightness of the object
- Dynamic range: The CCD has a wide linear dynamic range. That means we can observe bright and faint objects at the same time
- Low read out noise / dark current: Since we want to observe small signals we need a detector with a very low noise. This is fulfilled by the CCD. To acquire a low dark current we will cool down the CCD.
- Large wavelength spectrum: We need a detector sensitive in a large wavelength domain

We will now discuss how to get the best possible data from our CCD.

### 1.1 Bias and Dark Current

The ADC can only handle positive voltages. Therefore a bias is added by the electronics to the signal from the CCD. We will subtract this bias.

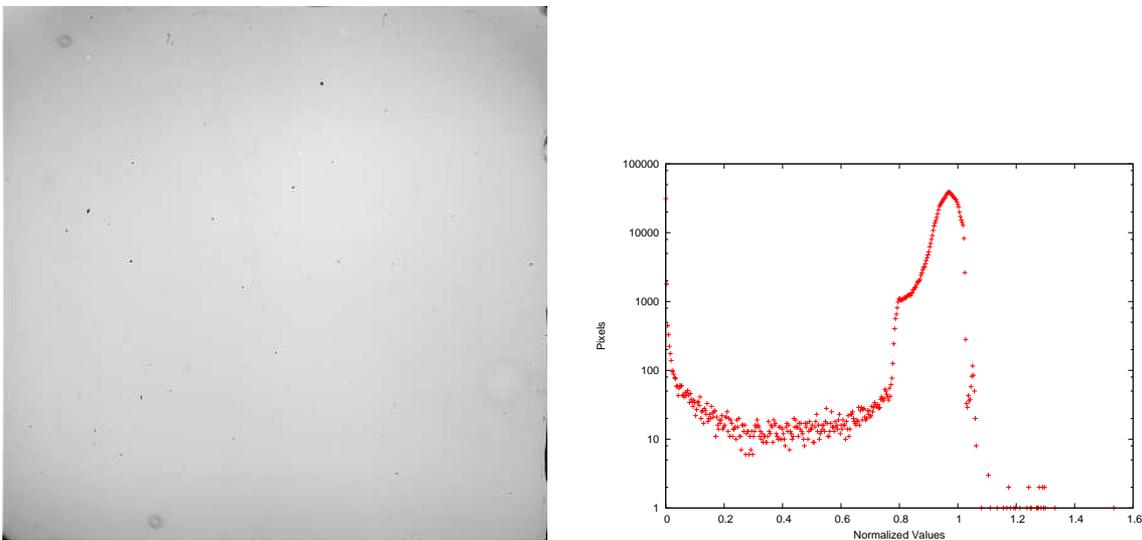
The dark current produces charges in the CCD because of thermal electrons. We will measure this charge although it does not originate from a photon hitting the CCD. Since the CCD is build of semi conductors, we will always have a current exponentially dependent on the temperature. We will deal with this by cooling the CCD detector down to roughly  $-115^{\circ}C$  using liquid nitrogen. To measure the dark current we will do exposures with the shutter closed, so we will not measure any photons hitting the detector but only the dark current. The dark current of course depends linearly on the exposure time. With decreasing temperature the dark current (average difference between zero and five seconds exposure time) went down below the noise, which consists of the read out noise and the signal noise. So we decide to ignore the dark current.

## 1.2 Flatfields

Due to irregularities during the construction process of the CCD not all pixels will be equally sensitive. We therefore will take a flatfield. This means we observe an uniformly illuminated source that is supposed to deliver us an identical signal for equally sensitive pixels. Though it does not. By dividing the flatfield by the average of the flatfield's pixels, we obtain an correction factor for each particular pixel. This is called master flatfield. By dividing an image by this master flatfield we correct pixel by pixel for the instrumental sensitivity.

Since the pixel sensitivity might be wavelength dependent, we need to create a master flatfield for each filter we will use in the experiment later. As illumination source for the flatfield either a dome flat can be used (equally illuminated white surface) or a sky flat (the sky just before the sun dusks but before the stars appear). For the sky flat the telescope needs to be dithered a bit and then median of several measurements is taken. Otherwise the flatfield will be disturbed by stars which the eye does not see but which are detected by the CCD.

Flatfield for V-Filter and histogram of number of pixels per correction value

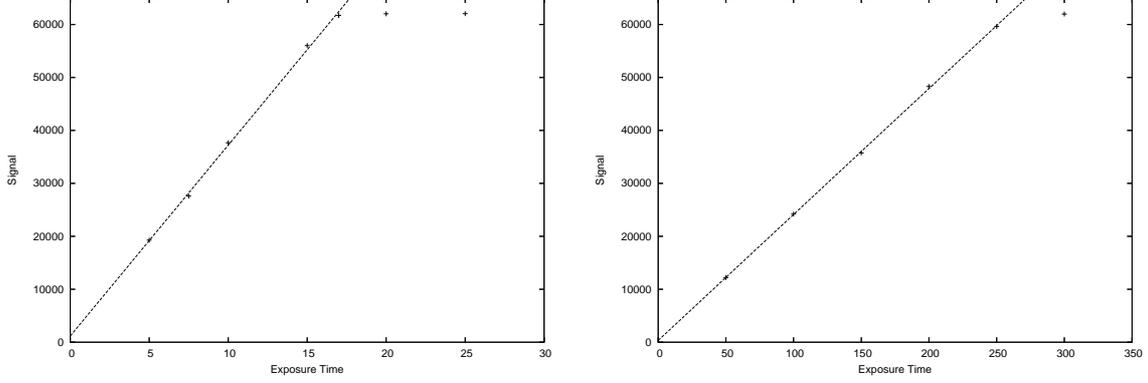


In the histogram we see, most pixels are in between of 0.75 and 1.05. (These are the variations we want to handle with the flatfields). Furthermore an increasing number of pixels with a value of zero are seen, measuring no noticeable signal. Those are dead pixel we will discuss later.

We will now use some flatfields to check some detector characteristics:

### 1.2.1 Linearity and dynamic range

We have taken flatfields for different exposure times. We now show plots of the average pixel values against the exposure time for the B and V Filter:



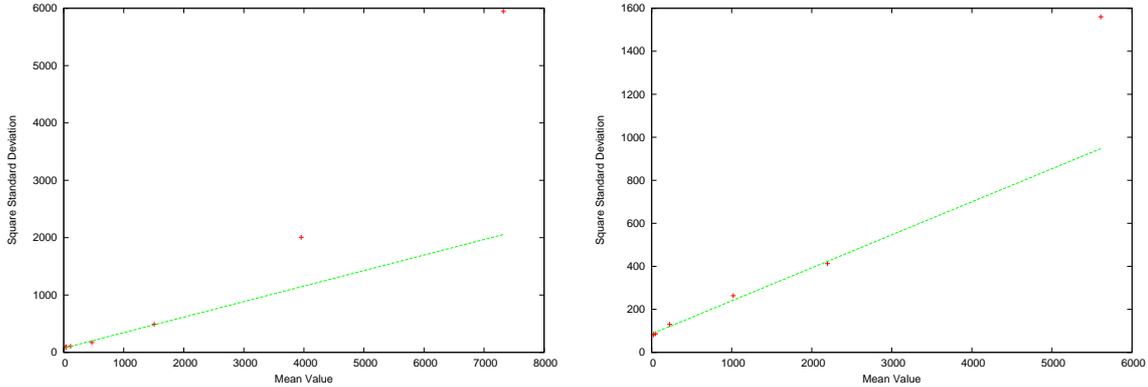
In both cases the dependency between signal and exposure time is linear and the detector saturates for about 62000 counts.

### 1.2.2 Read out noise and gain

The read out noise  $\sigma_R$  is produced by the read out electronics and is independent from other measurements. The gain  $\kappa$  is the number of excited electrons per unit of the ADU. The signal  $S$  is connected to the number of incoming photons  $N$  by  $\kappa S = \eta N$ , where  $\eta$  is the quantum efficiency. The variance of  $N$  according to Poisson statistics is  $\sigma_N = \sqrt{N}$ . So for  $\sigma_s$  we obtain:

$$\sigma_S^2 = \frac{S\eta}{\kappa} + \sigma_r^2$$

We now take flatfields for different illumination intensities and plot the variance of the signal we get from pixel each against the average signal itself. (We only take into account a center part of the image, where the image is almost uniform)



In both cases the expected linear dependency does not entirely appear. Thus to fit a line we only use the first four data points and obtain:

	V-Filter	B-Filter
$\sigma_R^2$	$73.82 \pm 11.01$	$86.08 \pm 7.01$
$\sigma_R$	$8.59 \pm 0.64$	$9.28 \pm 0.38$
$\eta/\kappa$	$0.271 \pm 0.0156$	$0.154 \pm 0.007$
$\kappa$	$3.32 \pm 0.19$	$4.87 \pm 0.22$

$$\eta_V = 90\% = 0.9 \quad \eta_B = 75\% = 0.75$$

Since the gain was set to 5 these values are compatible with the initial settings, we see especially for  $\kappa_V$  that our fit, as expected, was not good.

### 1.2.3 Dead Pixels / Cosmic Particles

It's colloquial to consider that some pixels of the CCD are damaged or dead and transmit no signal, while others may be hit by cosmic rays during the observation resulting in enough charge to saturate the pixel. Contrary to the dead pixels this will happen absolutely randomly. We will therefore repeat every measurement several times. This will eliminate the influence of cosmic rays, but for the dead pixels we need to dither the camera. This means we move the camera by a few arcsec, so light from a particular object, a star for example, will not hit the same dead pixel all the time. At the end we undo these shifts by manually shifting the images according to a reference image using reference stars and then combining the images by median. To combine the images using the median here is essential because we need to exclude very low and very high values to correct for dead pixels / cosmic rays.

### 1.2.4 Focus

Light passing the aperture is diffracted. The point spread function is the representation of a point source. For every filter we adjust the focus, so the full width at half maximum is minimal.

### 1.2.5 Photometry

There are two types of photometry to get an accurate magnitude measurement. Aperture photometry, which means integrating all counts within an aperture for every particular star. Afterwards the background has to be subtracted that can be obtained by averaging the counts a defined amount of pixels away from the center. This works well, if the stars on the image are not too close to each other. In a crowded field of stars however, the background cannot be obtained this way and stars might even overlap. There is another way of observation that can handle such crowded fields: so called PSF-Photometry. Here a point spread function is fitted to each star to obtain the magnitude. As we will observe an open cluster we will use aperture photometry.

## 2 Color-Magnitude diagram

### 2.1 Magnitude

First we define magnitude. We distinguish between different magnitude scales:

#### 2.1.1 Instrumental magnitude

This is, as the name says, the magnitude measured by a single instrument, and so can only be compared with values obtained with the same equipment. A logarithmic scale with an arbitrary zero point is used:  $m_{instrumental} = m_0 - 2.5 \cdot \log(S)$ .

#### 2.1.2 Apparent magnitude

Apparent magnitude is the brightness measured from Earth. This is needed to compare objects from different sources. A difference in brightness by a factor of 100 corresponds to a difference of 5 magnitudes. We use standard stars with known apparent magnitude to determine the offsets for the different filters. So for two stars this means:  $S_2 = S_1 \cdot 100^{(m_1 - m_2)/5}$  where  $S_1$  and  $S_2$  are the brightness,  $m_1$  and  $m_2$  the magnitudes of the stars.

### 2.1.3 Absolute magnitude

The absolute magnitude is defined as the apparent magnitude for objects standing at a distance of 10 pc from the sun. We know the measured signal is quadratically dependent on the distance, so for known apparent magnitude we can convert from absolute magnitude to distance and viceversa. For the apparent magnitude  $m$  and the absolute magnitude  $M$  this means:

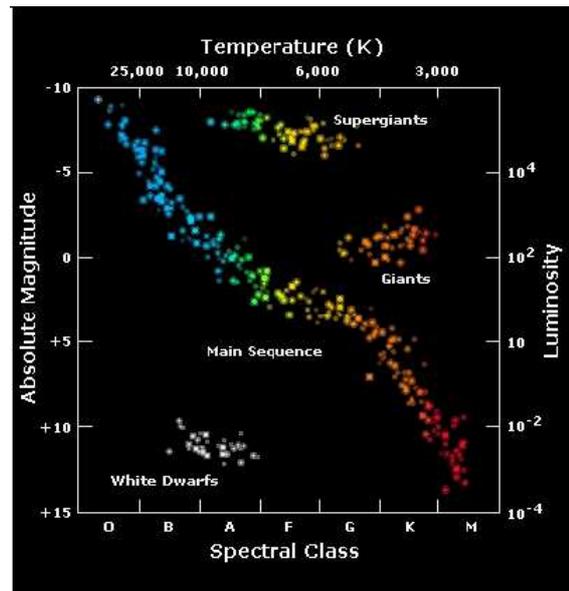
$$\frac{S_{10pc}}{S} = \left(\frac{d}{10 pc}\right)^2 = 100^{(m-M)/5}$$

The difference between absolute and apparent magnitude is called the distance modulus.

### 2.1.4 Hertzsprung-Russel-Diagram (HRD)

With known spectral-classes (color) and absolute magnitudes of many stars at the beginning of the 20th century Hertzsprung and Russel independently were able to do statistical analysis. The spectral class is the peak in the emitted frequency spectrum, that is, according to Planck, dependent on the surface temperature. The hotter and thus brighter and bluer stars on the left handed side of the diagram are bigger and heavier and burn their hydrogen, helium, etc. faster.

Obviously the absolute magnitude is not distributed independently of the spectral class. In the diagram stars appear in groups whereas most stars are in the main sequence.



### 2.1.5 Color-Magnitude-Diagram

As it is quite difficult to derive the spectral class and the absolute magnitude for stars to do a Hertzsprung-Russel-Diagram we will do a Color-Magnitude-Diagram. This means plotting the apparent magnitude for one filter against the color, that is the difference between the B and the V magnitude here. For the observation of objects like star clusters, as in our case, we can assume that all stars have approximately the same distance from Earth, so using the apparent instead of the absolute magnitude just shifts the diagram along the magnitude axis. The color B-V corresponds to the spectral class, so the diagram will look similar to the HRD.

### 2.1.6 Determining the distance

As said if we observe stars, that we assume to have the same distance, as in our case with open cluster M52, the color-magnitude-diagram is just shifted on the vertical magnitude scale. We compare the main sequence in our image with theoretical isochrones of star evolution or with the color-magnitude-diagram of a known star cluster with known magnitude. We then shift our diagram along the magnitude scale until the main sequences match. This shift is exactly the distance modulus we need to calculate the distance using the equation above.

### 2.1.7 Determining the age

We assume all stars within the cluster were formed at the same time. Since the brighter stars in the diagram have a shorter lifetime than fainter stars, they leave the main sequence earlier. So the main sequence will shift to fainter magnitude and redder color during the time. Using theoretical isochrones of star clusters with different ages we can compare the main sequence and so obtain the age of the cluster.

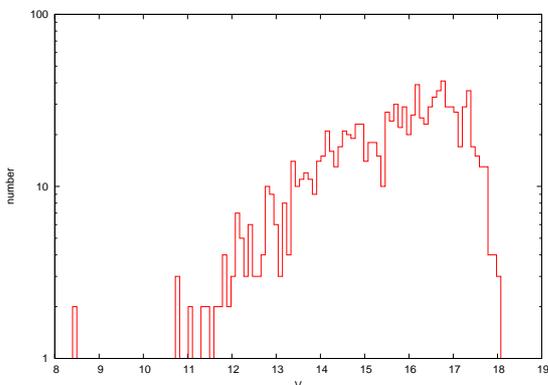
## 3 Experiment

### 3.1 Observation

We took 4 flatfields for both filters, B and V, that were not saturated. We subtracted the bias and combined them to the master flatfield. After adjusting the focus we took 6 images with the B filter and 7 with the V filter, while dithering the field of view by some arcsec between each exposure.

### 3.2 Evaluation

We subtracted the bias and applied the master flatfield correction to the scientific observations. Then we shifted the images manually to correct for the dithering and then combined the images to our final B and V images using the median. Then we used the IRAF package daofind to extract the stars from the images. We took only peaks that were 5 times the high of the background noise. This gave us 1629 stars for the B and 1491 stars for the V filter. We took 12 stars with known apparent magnitude and compared them with the magnitude we measured for them. We calculated the offset we need to get the apparent magnitude for all our stars using the median to 1.78 for the B and 1.32 for the V filter. We searched our lists for stars appearing in both list with the same coordinates and identified those stars. This gave us 1083 final stars with apparent B and V magnitude. We show an histogram of V magnitudes we measured:

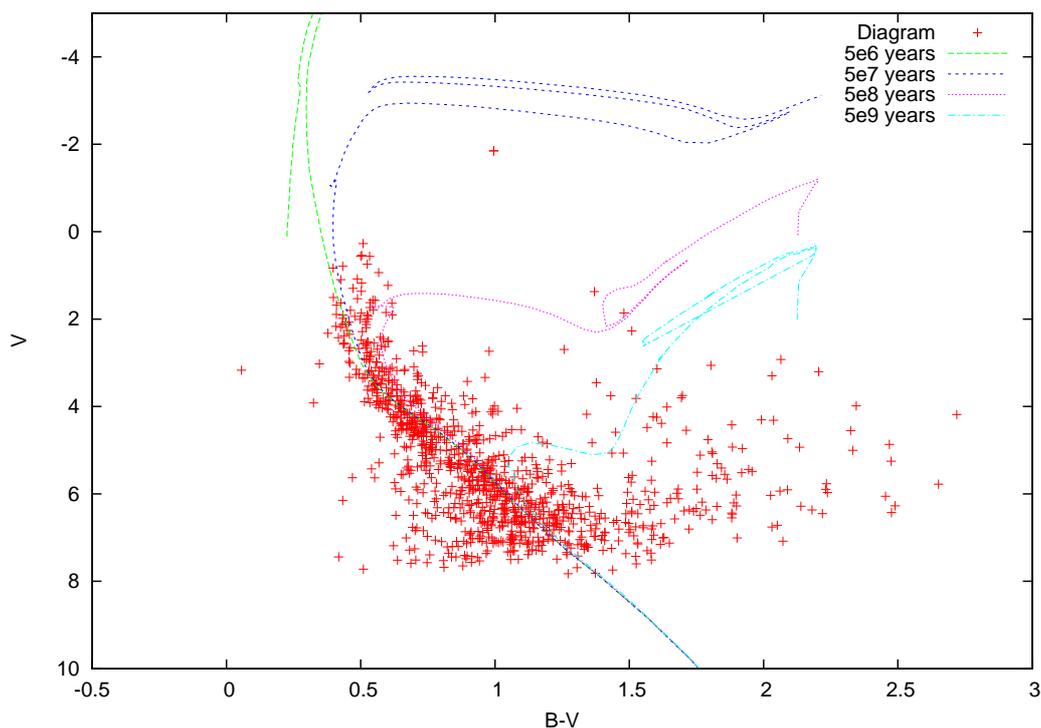


There are less brighter stars in the diagram (Brighter star correspond to lower magnitudes because of the inverse magnitude definition). This is just for the fact that there are less brighter than fainter stars present in the universe.

The faintest stars we measure, are at a magnitude of about 18. So our observation can go as deep as 18 in magnitude and we do not get those beyond 18. With bigger telescopes and increased observation time one would be able to reach deeper magnitudes.

### 3.3 Color-Magnitude-Diagram

We used those stars to determine the color-magnitude-diagram. In the diagram we also show isochrones for different ages. The main sequence is equal for all isochrones, but the turnoff point moves further to the left for older isochrones. Also we shifted our stars by 10.25 magnitudes along the V direction to best fit the observed main sequence to the isochrones. This shift of 10.25 is the distance modulus of our cluster M52. We will use this to calculate the distance.



The shift by 10.25 magnitudes corresponds to a distance of

$$d = 10^{(10.25+5)/5} \text{ [pc]} \approx 1122 \text{ [pc]}$$

The turnoff point of the main sequence corresponds best to our model for an age of about

$$5 \cdot 10^7 \text{ years}$$

Since we only compare with isochrones for an age of every power of 10, there is a large uncertainty in this value. The value though gives us a good hint to the order of magnitude of the age.

### 3.4 Comparison

To verify that our results are correct we searched sky catalogues for age and distance values of M52. Sky Catalogue 2000.0 gives an age of  $3.5 \cdot 10^7 \text{ years}$  what corresponds well to our age considering the logarithmic scale of our isochrones. The distance though varies from 920 to 2150 pc. So our value of 1122 pc lies within this range.

### 3.5 Conclusion

We cooled down the CCD, measured the dark current that we decided to ignore and then took dome flats to be ready to do our observation. Then we took images of the open cluster M52 using B and V Johnson filters. We subtracted the bias from our images and then divided them by the master flatfield. Then we did aperture photometry to get magnitudes of stars in the B and V band. Those we used to create a color-magnitude-diagram. We compared our diagram to theoretical isochrones and determined the distance modulus and the turnoff point of the main sequence. The distance modulus for M52 is 10.25 magnitudes which corresponds to a distance of about 1122 pc. The isochrone that best fit our main sequence was calculated for an age of  $5 \cdot 10^7 \text{ years}$ .