

THE WHAM SPECTROMETER:
DESIGN, PERFORMANCE CHARACTERISTICS,
& FIRST RESULTS

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Abstract

The Wisconsin H-Alpha Mapper (WHAM) is a unique new facility designed to study faint optical emission lines from diffuse ionized gas in the disk and halo of the Galaxy. WHAM consists of a 15 cm aperture, dual-etalon Fabry-Perot spectrometer coupled to a 0.6 m siderostat. Its primary purpose is to conduct a sensitive kinematic survey of the northern sky ($\delta > -30^\circ$) in the H α line from Kitt Peak with 1° spatial resolution and 12 km s^{-1} spectral resolution. This survey will provide the first detailed information about the warm ionized medium (WIM), a recently recognized major component of the interstellar medium.

After reviewing the basic properties of Fabry-Perot systems, I describe our specific technique for utilizing them, which involves the use of dual-etalons in series, pressure tuning, and imaging of the ring pattern passed by the etalons onto a CCD detector. A description of the WHAM optical system is also included. Measurements of the performance characteristics of the WHAM spectrometer relevant to producing a

calibrated spectrum are presented. I include detailed studies of the instrument profile function, linearity, wavelength and intensity calibration, the imaging performance of the optical system, sensitivity, the CCD detector properties, and noise. The spectrometer is found to meet or exceed its design specifications in all important aspects.

The first observations using the WHAM spectrometer include a study of emission lines from high velocity clouds in the M complex. The results include clear detections of $H\alpha$ emission from a number of locations toward the M I and M II clouds with intensities of $0.1 - 0.2 R$. Another study obtained the first detections of the He I $\lambda 5876$ recombination line from the WIM. This line has been detected in four directions in the Galactic plane, and the intensity ratios of the $\lambda 5876$ line to $H\alpha$ indicate that the spectrum of the radiation ionizing the WIM is significantly softer than that emitted by the O stars in the solar neighborhood.

To my family

“It is a miracle that curiosity survives formal education.”

Albert Einstein

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Chapter 1

Introduction

1.1 Why Study the Interstellar Medium?

The interstellar medium (ISM) plays a fundamental role in the process of Galactic evolution. This process involves a complex interplay between the interstellar matter and the stars. The stars are born from interstellar matter. During their lives, they deposit energy into the ISM in the form of electromagnetic radiation and stellar winds. When they die, the more massive stars return matter and energy back into the interstellar medium in dramatic supernova events. The matter returned to the ISM is enriched with heavy elements produced by nuclear burning in the stars' interiors and in processes occurring during the explosions. This enriched ejecta becomes the material of future generations of stars (which behave rather differently due to the presence of the “metals”). The process is not 100% efficient; some of the matter remains locked up forever in compact objects such as white dwarfs and neutron stars. Therefore, the composition of a galaxy slowly proceeds from all ISM to no ISM. The chemical evolution of a galaxy results from this cycle of stellar birth, death, and enriched rebirth.

The growth in quantity of these heavy elements allows for the possibility, along the way, of planets, rocks, and in at least one case living organisms.

The interstellar medium is arguably the least well understood player in this cycle. The main components of the interstellar medium have been identified, but the structure is still largely a mystery. The Galaxy is composed of primarily hydrogen (90% by number), some helium (10% by number), and only trace quantities of heavier elements. The gaseous component of the interstellar medium is known to contain regions of neutral atomic hydrogen (both in the form of clouds and in the form of a more diffuse component), molecular hydrogen, and ionized hydrogen (in the form of classical H II regions, a warm diffuse component, and a hot component). It is not known however, how these various components are distributed. For instance, it is still a matter of great controversy whether the hot component, thought to result from the ensemble of supernova remnants (SNR), fills most of the volume of the Galaxy, or is restricted to isolated SNR bubbles. Also, how the observed components fit together into an interacting physical system is poorly understood. One model (McKee & Ostriker 1977) contends that the diffuse ionized hydrogen in the ISM will be on the outer surfaces neutral hydrogen clouds. Other models (e.g. Miller & Cox 1993) speculate that the $H\alpha$ emitting gas is part of a more diffuse intercloud medium. Clearly, before the process by which galaxies evolve can be understood, a better understanding of the interstellar medium needs to be acquired.

Interstellar studies are also very important to all astronomers because virtually all astrophysical observations must look through the material in the ISM. The interstellar matter subtracts from (e.g., dust absorption, H I and He I absorption of UV radiation and soft X-rays), adds to (e.g., free-free emission from ionized gas, light

scattered into the beam by dust), and modifies (e.g., pulsar dispersions, Faraday rotation, polarization by aligned dust grains) the radiation on the way from the source to the observer. Detailed knowledge of the composition and distribution of interstellar matter is required to understand and remove the effects of this translucent screen. An obvious example is de-reddening of a stellar spectrum to correct for the bias introduced by interstellar dust. Another example, more closely related to the subject of this dissertation, is the removal of the free-free component of the microwave background using the WHAM survey of the diffuse ionized hydrogen.

1.2 The Warm Ionized Medium

The presence of ionized gas in the Galaxy has traditionally been associated with the bright ionized regions near hot stars called Strömgren spheres or classical H II regions. We now know that these classical H II regions contain only about 10% of the of the ionized hydrogen in the Galaxy. The remaining 90% is in the form of warm ($\sim 10^4$ K), low density ($\sim 0.1 \text{ cm}^{-3}$), fully ionized regions which fill approximately 20% of the volume within a 2 kpc thick layer about the Galactic plane. This warm ionized medium or WIM (sometimes also referred to as the Diffuse Ionized Gas or DIG), is now recognized as a major component of the interstellar medium.

The first evidence of widespread ionized hydrogen in the Galaxy came from peculiarities found in studies of the Galactic synchrotron background by Hoyle & Ellis (1963). More direct evidence of the presence of the diffuse component of H II came with the discovery of pulsars; the dispersion of the radio signals from these pulsars gives a direct measure of the column density of free electrons along the line of sight. Soon after this, $\text{H}\alpha$ emission from the WIM was detected using a large aperture Fabry-

Perot spectrometer (Reynolds, Roesler, & Scherb 1973).

The space averaged midplane density of the WIM was determined to be $\langle n_e \rangle_o \simeq 0.03 \text{ cm}^{-3}$ by determining the dispersion measure ($\text{DM} \equiv \int n_e dl$) toward a collection of pulsars with known distances. The discovery of pulsars in high latitude globular clusters provided the first measurement of the total column density of free electrons along high latitude sightlines. The result is that $N_{\text{HII}} = 0.8 - 1.4 \times 10^{20} \text{ cm}^{-2}$ from the midplane to the location of the globular clusters (Reynolds 1989a). Combining this with the midplane density gives the scale height $H \simeq N_{\text{HII}} / \langle n_e \rangle_o \simeq 1 \text{ kpc}$. Measuring the emission measure ($\text{EM} \equiv \int n_e^2 dl$) toward the same pulsar directions provided the first measurement of the volume filling fraction of the WIM with the result that $f_{\text{WIM}} \gtrsim 0.2$ (Reynolds 1991), implying that the diffuse ionized gas is clumped into regions with local densities $n_e \simeq 0.1 \text{ cm}^{-3}$. While the space averaged midplane density of ionized hydrogen is only 5% of that of the neutral hydrogen, the much larger scale height of the diffuse H II means that the total column density of ionized hydrogen through the disk of the galaxy is 1/4 to 1/2 that of the H I, and also implies that at $z \sim 1 \text{ kpc}$ the ionized gas may predominate.

The study of faint optical emission lines at high spectral resolution from the WIM has provided a great deal of valuable information about the distribution, kinematics, physical state, and energy requirements of this newly recognized major component of the ISM (see Table 1.1 for a summary of known properties). Fabry-Perot spectrometers are ideally suited to this type of study due to their ability to have high light throughput (high etendue) for diffuse sources at high spectral resolution (see Section 2.2.3). For example, Fabry-Perot measurements of the $\text{H}\alpha$ intensity at high Galactic latitudes imply a hydrogen recombination rate equal to 5×10^6 H-ionizations s^{-1} per cm^2 of

Galactic disk (Reynolds 1993). Each ionization costs 13.6 eV, and therefore, this high ionization rate leads to a substantial power requirement for the ionizing source (see below).

Faint optical emission lines also provide constraints on the temperature and ionization fraction in the WIM. A lower limit on the temperature $T_{WIM} \gtrsim 5400$ K results from measurements of the [O I] $\lambda 6300$ / $H\alpha$ ratio in combination with a measurement of the [S II] $\lambda 6716$ / $H\alpha$ ratio (Reynolds 1989b). An upper limit $T_{WIM} \lesssim 20,000$ K is derived from comparing line widths measured for the [S II] $\lambda 6716$ line to the widths of the corresponding $H\alpha$ line (Reynolds 1985). The generally adopted temperature is $T_{WIM} \simeq 8000$ K. A lower limit to the hydrogen ionization fraction χ_H within the ionized regions has been calculated from the upper limit [O I] $\lambda 6300$ / $H\alpha \leq 0.02$. This upper limit implies $\chi_H \geq 0.70$ for $T_{WIM} \geq 5400$ K, and if $T_{WIM} = 8000$ K, the lower limit becomes $\chi_H \geq 0.94$.

Another active area of research into the physical state of the WIM using faint optical emission lines is determining and understanding the helium ionization fraction χ_{He} . A recent measurement by Reynolds and Tufte (1995) of the He I $\lambda 5876$ recombination line from the WIM placed upper limits on the He I $\lambda 5876$ / $H\alpha$ ratio which lead to the conclusion that the helium in the WIM is significantly less ionized than the hydrogen. In Chapter 5, I present actual detections of this line at a level that is slightly higher than was suggested by the previous data.

The WIM also has distinctive spectral characteristics. Various emission lines from the WIM have been measured, and their strengths with respect to $H\alpha$ differ from the classical O star H II region line ratios. For example, the [S II] $\lambda 6716$ / $H\alpha$ ratios has been found to be much higher in the WIM than in the O star H II regions.

The $[\text{N II}] \lambda 6584 / \text{H}\alpha$ from the WIM is about the same as that from H II regions, while the $[\text{O III}] \lambda 5007 / \text{H}\alpha$ is low. Many of the spectral features found in the WIM can be explained in the context of an O star ionization model in which the stellar radiation field is very dilute. Models have been constructed by Mathis (1986) and by Domgörgen and Mathis (1994) attempting to explain these line ratios using O stars as the source of ionization. These models are fairly successful, but only work for certain distributions of the hydrogen in the ISM.

Since ionized hydrogen will naturally recombine to form neutral hydrogen, its presence requires an active source of ionization. The average recombination rate per cm^2 of Galactic disk is $\langle r_G \rangle \simeq 5 \times 10^6 \text{ s}^{-1} \text{ cm}^{-2}$, which sets the required ionizing photon production rate, or at 13.6 eV per ionization a power requirement of $\langle W \rangle \simeq 1.0 \times 10^{-4} \text{ ergs s}^{-1}$ per cm^2 of Galactic disk. Of the known sources of ionization, only the Lyman continuum flux of O stars comfortably exceeds this energy requirement. Other possible sources include supernovae (100% of the energy injected by supernova is required to maintain the ionization), cosmic rays, turbulent mixing layers, Galactic flares, etc...

There are several difficulties associated with the idea of O stars ionizing the interstellar medium. The first is a spatial problem. The O stars reside in the Galactic plane and are few and far between. The WIM is a very thick layer and is in large part far from the O stars. It takes very little H I gas to absorb the Lyman continuum flux from the O stars on its way to the WIM gas and there is considerable H I in the Galaxy. Hence, a very special arrangement of the H I is necessary to match simultaneously the distribution of column density of H I on the sky, known from 21 cm maps, and the constraint that the O star photons can reach the WIM. Models have been constructed

by Miller & Cox (1993) and by Dove & Shull (1994) to investigate the feasibility of this. Both find that it can be done if the H I distribution satisfies certain constraints, constraints which may or may not exist in the real ISM. The other problem with the O star ionization model is a spectral problem. The He I $\lambda 5876$ / $H\alpha$ line has been measured to be quite low compared to classical H II regions. This indicates a softer spectrum than is known to be emitted by the ensemble of O stars in the solar neighborhood (see Chapter 5).

In summary, although the average properties of the WIM are now known, detailed information is not yet available about the distribution and relation to the H I, ionization state, temperature, and relationship to known sources of ionization such as O stars.

1.3 The Wisconsin H-Alpha Mapper (WHAM)

The WHAM facility is a recently completed dual-etalon Fabry-Perot spectrometer coupled to a 0.6 m siderostat designed to study faint optical emission lines from diffuse sources at high spectral resolution. The WHAM system is 100 times more sensitive than the previously most sensitive instrument of its type, and therefore opens up new opportunities for exploration. The primary purpose of the WHAM instrument is to conduct a sensitive kinematic survey of the northern sky ($\delta > -30^\circ$) in the Balmer- α line of hydrogen ($H\alpha$), analogous to the surveys of H I in 21 cm. The survey will have 1° spatial resolution and 12 km s^{-1} spectral resolution. The 30 second integration time per beam will provide a signal to noise ratio of 20 for a 0.5 R line having a width of 20 km s^{-1} (FWHM). The WHAM instrument is now operating at its Kitt Peak site, and the survey will begin in January of 1997. We expect the approximately 40,000

spectra will take about 2 years to gather, and that the calibrated spectra and velocity interval maps will be available within a year after the data is collected. The WHAM $H\alpha$ survey will provide the first detailed look at the distribution and kinematics of the ionized gas in the Galaxy. This data set will also be compared to already existing H I surveys, which will allow study of the detailed inter-relationship between the neutral and ionized gas in the interstellar medium.

In the first part of this dissertation, I present a description of the design of the WHAM spectrometer, and measurements of the performance characteristics. In Chapter 2, after reviewing the basic theory of Fabry-Perot systems, I describe our specific technique for utilizing them, which involves the use of dual-etalons in series, pressure tuning, and imaging of the ring pattern passed by the etalons onto a CCD detector. A description of the optical system is also included. Measurements of the performance characteristics relevant to producing a calibrated spectrum with the WHAM spectrometer are presented in Chapter 3. I include studies of the instrument profile function, linearity, wavelength and intensity calibrations, the imaging performance of the optical system, sensitivity, the CCD detector properties, and noise.

By using long exposures (e.g., 900 seconds), the WHAM instrument is capable of studying very faint ($\lesssim 0.1$ R) optical emission lines. This capability was exploited while the instrument was located at its Pine Bluff Observatory site west of Madison (January 1996 – November 1996). In one study, we observed $H\alpha$ emission from high velocity clouds (HVCs) in the M complex. The results include a number of first detections of such emission and are presented in Chapter 4 of this dissertation. Another study obtained the first detections of the He I $\lambda 5876$ recombination line from the WIM. The intensity of this line compared to the $H\alpha$ intensity probes the degree of helium

ionization in the WIM, which is related to the hardness of the ionizing spectrum. The results of this study are presented in Chapter 5.

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	Parameter	Value
Geometry	Scale height, H_e	910 (-320, +400) pc
	Average volume filling fraction, $\langle f_e \rangle$	$> 20\%$
	Average local density, n_e	$0.08^a \pm 0.02 \text{ cm}^{-3}$
	Midplane local density ^b , $(n_e)_{z=0}$	$\sim 0.16 \text{ cm}^{-3}$
	Surface density	$1.4 \pm 0.2 \times 10^{20} \text{ cm}^{-2}$ $1.6^c \text{ M}_\odot \text{ pc}^{-2}$
Energy	Temperature, T_e	$\sim 8000 \text{ K}$
	Non-thermal speeds	$12\text{--}30 \text{ km s}^{-1}$
	Ionization rate, $\langle r_G \rangle$	$4.8 \pm 0.4 \times 10^6 \text{ s}^{-1} \text{ cm}^{-2}$
Ionization	H ⁺ fraction, $n(\text{H}^+)/n(\text{H}^0)$	> 2 (if $T_e \sim 5400 \text{ K}$) > 15 (if $T_e \sim 8000 \text{ K}$)
	He ⁺ fraction, $n(\text{He}^+)/n(\text{He}^0)$	$\lesssim 0.27$

Table 1.1: Properties of the Warm Ionized Medium

^aWithin $z \sim \pm 1 \text{ kpc}$.

^bAssuming an exponential distribution and H_e above.

^cH + He.

Chapter 2

The WHAM Spectrometer: Theory and Design

2.1 Introduction

The Wisconsin H-Alpha Mapper (WHAM) is a unique new instrument for the detection of faint emission lines from diffuse ionized gas in the disk and halo of the Galaxy. WHAM consists of a 15 cm aperture, dual-etalon Fabry-Perot spectrometer coupled to a 0.6 m siderostat. This provides a one degree diameter beam on the sky and produces a 12 km s^{-1} resolution spectrum over a 200 km s^{-1} interval that can be selected at any wavelength within the range $4800 \text{ \AA}$ to $7200 \text{ \AA}$. In this chapter I present the basic theory behind Fabry-Perot systems and describe how the WHAM spectrometer was designed to exploit the properties of the Fabry-Perot for the purpose of high-resolution spectroscopy of emission lines from diffuse ionized gas. I include a description of our method of employing Fabry-Perots, which involves the use of dual-etalons in series, pressure tuning, and imaging of the ring pattern passed by the etalons

onto a CCD detector.

2.2 Introduction to Fabry-Perot Spectroscopy

In this section the basic theory behind Fabry-Perot systems is described. I first derive the wavelength dependent transmission function of the Fabry-Perot and then discuss some of the important spectroscopic properties of this device.

2.2.1 The Transmission Function of a Fabry-Perot

The heart of a Fabry-Perot system consists of two accurately parallel reflecting surfaces with some spacing, l , between them. This is often accomplished by depositing a reflective coating on one side of two disks of fused silica, and then separating the reflecting surfaces by spacers of precisely identical length. Such an arrangement is referred to as a Fabry-Perot ‘etalon’. Imagine that this system is illuminated from one side as in Figure 2.2. The basic idea is that multiple reflections occur between the reflective surfaces and the interference between these reflections gives rise to the unique wavelength dependent reflection and transmission properties of this device.

A brief digression is in order here to derive the expression for the the phase lag, ϕ , due to the optical path difference between two consecutive transmitted rays. Figure 2.1 shows an etalon system with an index of refraction n_g in the gap, immersed in a medium with index n , and illuminated from the left. Also depicted is a ray representing the portion of the incoming wave that is transmitted directly (labeled 1), and one that represents the portion of the incoming ray that is transmitted after two internal reflections (labeled 2). Further internal reflections result in an identical optical path difference between consecutive transmitted rays.

Let us derive the phase lag between rays 1 and 2 in Figure 2.1. In general,

$$\phi = \frac{2\pi}{\lambda} \Delta L_{opt}, \quad (2.1)$$

where ΔL_{opt} is the optical path difference between the two rays. From Figure 2.1 it is apparent that

$$\Delta L_{opt} = n_g(AB + BC) - n(AD).$$

But,

$$(AB + BC) = \frac{2l}{\cos \theta_t}$$

and,

$$AD = (2l \tan \theta_t)(\sin \theta_i).$$

Hence,

$$\Delta L_{opt} = n_g \frac{2l}{\cos \theta_t} - n(2l \tan \theta_t)(\sin \theta_i).$$

Using Snell's law, $n \sin \theta_i = n_g \sin \theta_t$, we have

$$\begin{aligned} \Delta L_{opt} &= 2ln_g \left(\frac{1}{\cos \theta_t} - \frac{\sin^2 \theta_t}{\cos \theta_t} \right) \\ &= 2n_g l \cos \theta_t. \end{aligned}$$

Then, using Equation 2.1, we have

$$\phi = \frac{4\pi n_g l}{\lambda} \cos \theta_t. \quad (2.2)$$

Let us now derive the transmission function, I_t , for the simple configuration drawn in Figure 2.2, where we have two surfaces separated by a gap of length l and index of refraction n_g immersed in a material of index n . We begin by calculating the amplitude transmission function, a_t . A plane wave of amplitude $a_o \exp[i(\omega\tau + kx + \alpha)]$

is incident at an angle θ_i to the surface normal, where α is the phase advancement necessary to make the phase of the directly transmitted ray equal to zero at the plane where the transmitted intensity will be evaluated (represented by the dotted line in Figure 2.2). Let r be the amplitude fraction that is reflected at the surface, and t be the amplitude fraction that is transmitted. Note that τ is used here to represent the time to avoid confusion with the transmission coefficient, t .

The amplitude of the wave that is transmitted directly is $a_o t^2 \exp(i\omega\tau)$. After two internal reflections a wave of amplitude $a_o t^2 r^2 \exp(i(\omega\tau - \phi))$ is also transmitted. Similarly, there is a wave that suffers four internal reflections with amplitude $a_o t^2 r^4 \exp\{i(\omega\tau - 2\phi)\}$ that is transmitted. In fact, there are an infinite number of successively weaker reflections that are transmitted, and it is the sum of these interfering waves that determine the amplitude of the transmitted wave. To wit, if we call the transmitted wave a_t ,

$$\begin{aligned} a_t &= a_o t^2 \exp(i\omega\tau) + a_o t^2 r^2 \exp[i(\omega\tau - \phi)] + a_o t^2 r^4 \exp[i(\omega\tau - 2\phi)] + \cdots \\ &= a_o t^2 \exp(i\omega\tau) [1 + r^2 \exp(-i\phi) + r^4 \exp(-2i\phi) + \cdots]. \end{aligned}$$

Let $x \equiv r^2 \exp(-i\phi)$. Then,

$$a_t = a_o t^2 \exp(i\omega\tau) \sum_{j=0}^{\infty} x^j.$$

But,

$$\sum_{j=0}^{\infty} x^j = \frac{1}{1 - x}.$$

Hence,

$$a_t = \frac{a_o t^2 \exp(i\omega\tau)}{1 - r^2 \exp(-i\phi)}.$$

To get the transmitted intensity, I_t , we must multiply a_t by its complex conjugate,

$$\begin{aligned} I_t &= a_t a_t^* \\ &= \frac{a_o^2 t^4}{1 - r^2 \{\exp(i\phi) + \exp(-i\phi)\} + r^4} \\ &= \frac{a_o^2 t^4}{1 - 2r^2 \cos \phi + r^4}. \end{aligned}$$

Therefore, with the identity $\cos \phi = (1 - 2 \sin^2 \frac{\phi}{2})$, we have

$$\begin{aligned} I_t &= \frac{a_o t^4}{1 - 2r^2 + 4r^2 \sin^2 \frac{\phi}{2} + r^4} \\ &= \frac{a_o t^4}{(1 - r^2)^2 + 4r^2 \sin^2 \frac{\phi}{2}} \\ &= \frac{a_o t^4}{(1 - r^2)^2} \frac{1}{(1 + \frac{4r^2}{(1-r^2)^2} \sin^2 \frac{\phi}{2})}. \end{aligned}$$

Define $R \equiv r^2$, $T \equiv t^2$, and A such that $T + R + A = 1$. These are called the reflectance, transmittance, and absorbance, respectively, of the reflecting surface. Then,

$$\frac{I_t}{I_o} = \frac{T^2}{(1 - R)^2} \mathcal{A}(\phi), \quad (2.3)$$

where I_o is the incoming intensity and where we have made use of the definition of the Airy function, $\mathcal{A}(\phi)$,

$$\mathcal{A}(\phi) \equiv \frac{1}{1 + \frac{4R}{(1-R)^2} \sin^2 \frac{\phi}{2}}. \quad (2.4)$$

This may also be written

$$\frac{I_t}{I_o} = \left(1 - \frac{A}{(1 - R)}\right)^2 \mathcal{A}(\phi). \quad (2.5)$$

In the special case where $A = 0$ and thus $T = 1 - R$, we have

$$\frac{I_t}{I_o} = \mathcal{A}(\phi). \quad (2.6)$$

A very similar analysis can be performed to derive the wavelength dependence of the *reflected* light. The main difference is that the first reflection is π out of phase with all the subsequent reflections because it is an external reflection. The result, in the case of no absorption, is

$$\frac{I_r}{I_o} = \frac{\frac{4R}{(1-R)^2} \sin^2 \frac{\phi}{2}}{1 + \frac{4R}{(1-R)^2} \sin^2 \frac{\phi}{2}}. \quad (2.7)$$

As I will show below, it is desirable to make R fairly high for good resolution and contrast, and Equation 2.5 shows that good peak transmission therefore requires A to be quite low. It wasn't possible to make a reflecting surface with both high R and low A until the development of techniques for fabricating dielectric multi-layer coatings (circa 1950). This explains why it is that even though the Fabry-Perot concept is quite old (one was constructed by Charles Fabry and Alfred Perot in the late 1800s), they have not been used extensively until fairly recently.

2.2.2 Some Important Spectroscopic Properties of the Fabry-Perot

Figure 2.3 shows a graph of the wavelength dependent transmission function for an ideal Fabry-Perot system with $R = 0.90$, and Figure 2.4 shows the reflection function. The peaks in the transmission function occur when $\sin^2(\phi/2)$ is at a minimum, which is at $\phi = 0, 2\pi, 4\pi, \dots$. Thus, the criterion for maximum transmission is

$$2n_g l \cos \theta_t = m\lambda, \text{ where } m = 0, 1, 2, \dots \quad (2.8)$$

It is often convenient to express things in terms of the reciprocal wavelength, $\sigma = 1/\lambda$. Our last equation then becomes

$$2n_g l \cos \theta_t \sigma = m \text{ where } m = 0, 1, 2, \dots \quad (2.9)$$

As is apparent in Figure 2.3 and in Equation 2.8, the Fabry-Perot transmission is a periodic function, and for a given n_g , l , and $\cos \theta_t$, there will be multiple orders passed. Different orders are labeled by the index m and correspond to different numbers of wavelengths fitting into the path difference. Since the Fabry-Perot spectroscopist will be counting photons in one way or another, and cannot distinguish between photons from different orders, it is usually necessary to block all but a single order. An important parameter in a Fabry-Perot is the distance to the next order, which is called the free spectral range (it is the spectral range that is free of confusion from other orders) and is denoted Q (see Figure 2.3). A subscript is used to indicate which parameter is being considered, so for instance, Q_ϕ is the amount by which ϕ needs to change to reach the next order. From Equation 2.4 it is clear that $Q_\phi = 2\pi$. It is also interesting to know Q_σ . Consider two adjacent orders:

$$2n_g l \cos \theta_t \sigma_1 = m_1, \quad \text{and}$$

$$2n_g l \cos \theta_t \sigma_2 = m_2.$$

Subtracting these two equations gives

$$2n_g l \cos \theta_t (\sigma_1 - \sigma_2) = m_1 - m_2.$$

But, we assumed the orders were adjacent implying $m_1 - m_2 = 1$, and hence

$$\sigma_1 - \sigma_2 = \frac{1}{2n_g l \cos \theta_t}.$$

Therefore, on axis, where $\cos \theta_t \simeq 1$, we have

$$Q_\sigma = \frac{1}{2n_g l}. \tag{2.10}$$

As I will discuss below, one of the main reasons to employ two etalons in series is to extend the free spectral range of the system.

The spectral resolution achievable with a Fabry-Perot is simply related to the width of the transmission peaks shown in Figure 2.3. Let $(\Delta\phi)_{FWHM}$ denote the full width at half maximum in ϕ of the transmission function. In the ideal Fabry-Perot, the half maximum occurs where

$$\frac{1}{1 + \frac{4R}{(1-R)^2} \sin^2 \frac{\phi_{1/2}}{2}} = 1/2,$$

where $\phi_{1/2}$ is the value of ϕ at which the transmission function is 0.5 times the peak value. Using $\sin(\phi_{1/2}/2) \simeq \phi_{1/2}/2$, we have

$$\frac{4R}{(1-R)^2} \frac{\phi_{1/2}^2}{4} = 1,$$

and thus

$$\phi_{1/2} = \frac{(1-R)}{\sqrt{R}}.$$

But, the full width at half maximum is just twice $\phi_{1/2}$, so

$$(\Delta\phi)_{FWHM} = \frac{2(1-R)}{\sqrt{R}}. \quad (2.11)$$

A number that is commonly used to characterize a Fabry-Perot system is the number of spectral elements in a free spectral range. This is referred to as the reflective finesse (or often just the finesse), $\mathcal{F}$, where

$$\mathcal{F} = Q_\phi / (\Delta\phi)_{FWHM}. \quad (2.12)$$

Using $Q_\phi = 2\pi$ and Equation 2.11, we have

$$\mathcal{F} = \frac{\pi\sqrt{R}}{(1-R)}. \quad (2.13)$$

Figure 2.5 shows a graph of the finesse versus R .

Another important property of the Fabry-Perot (or for any spectrometer for that matter) is the resolving power, $\mathcal{R}$, defined by

$$\mathcal{R} \equiv \frac{\lambda}{(\Delta\lambda)_{min}}, \quad (2.14)$$

where $(\Delta\lambda)_{min}$ is the smallest resolvable wavelength difference. There are various specific ways of defining $(\Delta\lambda)_{min}$, but for convenience we will take it to be the full width at half maximum (FWHM) of the transmission function. We call the resolving power for the ideal Fabry-Perot the ‘theoretical resolving power’, $\mathcal{R}_o$. The resolution of a real Fabry-Perot system will be less than $\mathcal{R}_o$ due to effects such as non-flatness of the plates (see Sections 2.3.1 and 2.3.6). From Equation 2.2 it is clear that

$$\frac{\phi}{\Delta\phi} = \frac{\sigma}{\Delta\sigma} = \frac{\lambda}{\Delta\lambda}. \quad (2.15)$$

Using Equation 2.2 and Equation 2.11, we have

$$\mathcal{R}_o = \frac{\phi}{\Delta\phi} = \frac{\pi\sqrt{R}}{(1-R)} 2n_g l \cos\theta_t \sigma, \quad (2.16)$$

or, more compactly,

$$\mathcal{R}_o = \mathcal{F}m, \quad (2.17)$$

where m is the order number.

Another important spectroscopic property is the contrast, C , the ratio of the maximum transmission to the minimum transmission of the spectrometer,

$$C = \frac{I_t(max)}{I_t(min)}. \quad (2.18)$$

I will derive the contrast for the case of a single etalon with no absorption. The transmission maxima occur where the $\sin^2(\phi/2)$ term in Equation 2.6 is zero,

$$I_t(max) = I_o. \quad (2.19)$$

This is an amazing thing: one can shine light onto two 90% reflecting surfaces and obtain nearly 100% transmission (but only for certain wavelengths). The minimum occurs where the $\sin^2(\phi/2)$ term is unity,

$$I_t(min) = \frac{I_o}{1 + \frac{4R}{(1-R)^2}}. \quad (2.20)$$

And hence, the contrast for the ideal Fabry-Perot with no absorption is

$$C = 1 + \frac{4R}{(1-R)^2} = \frac{(1+R)^2}{(1-R)^2}, \quad (2.21)$$

or, in terms of the finesse,

$$C = 1 + \frac{4\mathcal{F}^2}{\pi^2}. \quad (2.22)$$

In a real Fabry-Perot system, the contrast will be somewhat degraded due to scattered light. The contrast is important to us for two reasons: 1) It determines how much continuum light gets through off line and thus affects the background level and hence the noise, and 2) it determines the strength of the ghost peaks in the transmission function which are passed by one etalon but rejected by the other (see below).

2.2.3 The Throughput Advantage of Fabry-Perots Over Grating Spectrometers for Diffuse Sources of Emission

The main advantage of using a Fabry-Perot spectrometer instead of a diffraction grating based spectrometer is that the Fabry-Perot is capable of a much higher throughput at high spectral resolution than is a grating system. This advantage arises from the ability of the Fabry-Perot system to pass a large solid angle, Ω , of light into a small spectral element size, $d\lambda$. Consider the basic relations governing the two types of spectrometers:

$$\text{Fabry Perot: } \lambda \propto \cos \theta$$

Diffraction Grating: $\lambda \propto \sin \theta$

The crux of the issue lies in the difference between the behavior of the $\cos \theta$ function and the $\sin \theta$ function for small θ . For a given wavelength interval, $d\lambda$, the magnitude of the opening angle $d\theta$ is given by:

$$\text{Fabry-Perot: } d\theta \propto \frac{1}{\sin \theta} d\lambda$$

$$\text{Grating: } d\theta \propto \frac{1}{\cos \theta} d\lambda.$$

Hence, for a given $d\lambda$ a much larger $d\theta$ is allowed in the case of the Fabry-Perot.

This throughput advantage is only realized if the source of emission is extended and thus able to fill the large $d\theta$ allowed by the Fabry-Perot. For a star, the solid angle passed by the system is usually limited by the source and not the spectrometer, and thus the diffraction grating is no worse off than the Fabry-Perot. But for diffuse emission from the ISM, the solid angle advantage of the Fabry-Perot can be enormous; gains of more than 100 times over a diffraction grating operated at the same spectral resolution are not atypical.

2.3 Using Fabry-Perots for High-Resolution Interstellar Spectroscopy

The ultimate goal of the spectroscopist is to measure the intensity of a source as a function of wavelength. In this section I will examine how we have exploited the properties of the Fabry-Perot described above to construct a spectrometer optimized for high-resolution spectroscopy of faint optical emission lines from the diffuse interstellar medium.

2.3.1 Choosing and Setting the Resolution

The desired resolution of a spectrometer is determined by the characteristics of the source under study and the objectives of the study. The WHAM spectrometer is designed to study emission lines originating in the interstellar medium. The natural width of these lines is very small, and the profile shape arises from the motions of the gas atoms, both thermal motions and bulk motions, through the Doppler effect. For this reason we often consider the abscissa of our spectra a velocity scale instead of a wavelength scale. Emission lines from the warm ionized component of the interstellar medium typically have widths of $15 - 50 \text{ km s}^{-1}$, and the bulk motions of the gas are usually within $\pm 100 \text{ km s}^{-1}$ of the earth's velocity. A resolution of 12 km s^{-1} is adequate to separate the interstellar emission from any terrestrial emission and thus probe the kinematics of the emitting gas, as well as to probe line-broadening mechanisms, both thermal and non-thermal, through determinations of the line widths. We set our theoretical resolving power to achieve a velocity resolution of 8 km s^{-1} in order to allow for other instrumental broadening effects described below. The wavelength change due to the Doppler shift is governed by the formula

$$\frac{\Delta\lambda}{\lambda} = \frac{v}{c}, \quad (2.23)$$

where v is the velocity of the emitter with respect to the observer, and c is the velocity of light. Hence, our required velocity resolution of 8 km s^{-1} leads to a resolving power, $\mathcal{R}_o = 37500$.

From Equation 2.16, we see that the two ways of adjusting the resolution of the Fabry-Perot etalon are through the reflectivity of the etalon surfaces, R , and through the gap spacing between the plates, l . Since increasing l also decreases the free spectral

range, it would seem preferable to increase R to reach high resolution. However, the tolerance for errors in the flatness of the plates, and for nonparallelness of the plates decreases with increasing R . Therefore, the design strategy for reaching high resolution is to acquire the flattest plates that can be afforded (or that can be manufactured), set R as high as possible without letting the resolution be degraded by the non-flatness of the plates, and then set l to achieve the desired resolving power. What non-flatness can be tolerated for a given R ? For on-axis rays, a change in the gap spacing, l , by $\lambda/2$ moves the transmitted wavelength to the next order, that is, $Q_l = 1/(2\lambda)$. Hence, changing l by $Q_l/\mathcal{F}$ moves the transmission peak by a half-width. Therefore, the tolerance for plate defects and non-parallelness in the plates is

$$(\Delta l)_{\text{defects}} < \frac{\lambda}{2\mathcal{F}}. \quad (2.24)$$

The plates used in the WHAM spectrometer are flat to around $\lambda/150$, implying that $\mathcal{F}$ should be less than 75, implying $R < 0.96$. The surfaces of the WHAM etalons are coated such that $R = 0.90$, hence $\mathcal{F} = 29.8$ and thus a plate defect of $\lambda / 60$ will lead to a degradation of the resolution in our system. Hence, with our plate flatness we do not expect such a degradation.

With the finesse $\mathcal{F} = 29.8$, and a desired resolving power $\mathcal{R}_o = 37500$, Equation 2.17 determines the order, $m = 1258$. Hence, near $H\alpha$, $l = (6563\text{\AA})(1258.5)/2 = 0.041$. In fact, the gap in the high-resolution etalon of the WHAM spectrometer is $l = 0.0471$ cm, and the theoretical resolving power is $\mathcal{R} \simeq 43000$.

2.3.2 The Dual-etalon Concept

The primary purpose of placing two Fabry-Perot etalons in series is to increase the free spectral range of the system. One can also increase the free spectral range by

decreasing the gap spacing l , but that simultaneously degrades the spectral resolution. There are two other added advantages in that the product of the second Airy function with the first both reduces the wide, Lorentzian-like wings of the primary Airy function and greatly increases the contrast of the system.

For the WHAM spectrometer the primary, higher resolution etalon is combined with a lower resolution etalon. The transmission peaks of the two etalons can then be lined up for the desired wavelength (see Section 2.3.3 below), and the different free spectral ranges of the two etalons are such that only after a large number of free spectral ranges are the peaks lined up again. Figure 2.6 shows the transmission function of each of the WHAM etalons along with the combined transmission function, which is the product of the individual transmission functions. The high resolution etalon has a gap spacing of 0.0471 cm, and the lower resolution etalon has a spacing of 0.0201cm, leading to $Q_{\sigma}(\text{high}) = 10.62 \text{ cm}^{-1}$ and $Q_{\sigma}(\text{low}) = 24.88 \text{ cm}^{-1}$. It can be seen in Figure 2.6c that the combined transmission function, which is shown here with the two etalons tuned to $\text{H}\alpha$, has a single strong transmission peak along with some very weak 'ghost' peaks that appear where only one etalon has a transmission peak. These ghost peaks in the transmission function are down by of order the contrast, which for $R = 0.90$ is $C \simeq 360$. Thus we expect the ghost peaks to be of order 1/3 percent. This is confirmed in Figure 2.6d where the combined transmission function is plotted on a log scale. Where neither etalon has a transmission peak, the dual-etalon transmission function goes to a very low level, and the contrast of the combined system is of order the product of the contrasts of the individual systems.

It is still necessary to use an interference filter as a pre-monochrometer since the transmission peaks of the two etalons will eventually line up again, but the filter can

have a much broader pass band than would be necessary for a single etalon system. The filter is also useful in suppressing the parasitic light, that is, continuum light leaking through away from the transmission peak due to the finite contrast. A typical transmission function for one of the WHAM interference filters (20 Å FWHM) is shown as a dashed line in Figure 2.6a.

An important point to note about the dual-etalon transmission function relates to the $\cos \theta_t$ dependence in Equation 2.8. The transmission functions plotted in Figure 2.6 correspond to on-axis rays, where $\cos \theta_t = 1$. For rays at an angle θ_t with respect to the optical axis, the entire pattern of the transmission peaks is shifted to shorter wavelengths. Fortunately for Fabry-Perots, the angular dependence of the transmitted wavelength is independent of the resolution of the etalon, and thus the two etalons maintain the same relationship for off-axis rays.

The Airy function is not a Gaussian: it has significant wings, allowing wavelengths fairly far from the central peak to leak through at a low level. Figure 2.7 shows how the second, lower resolution etalon serves to suppress these wings. Figure 2.7a shows the transmission function of the high resolution etalon alone and allows a detailed look at the shape of the Airy function, and Figure 2.7b shows the low resolution transmission function. Figure 2.7c shows the product transmission function as a solid line and shows the profile of the high resolution etalon alone overlayed as a dashed line in order to demonstrate the wing suppression provided by dual etalons.

2.3.3 Pressure Tuning

In order to successfully implement a system with two Fabry-Perot etalons in series, it is necessary to be able to align transmission peaks of the two etalons at a

desired wavelength. This is called *tuning* the system. In our case, this is accomplished by making slight changes to the optical gap, nl , in each etalon through the dependence of the index of refraction n of the gas in the gap, on pressure, p . Each etalon is immersed in a separate chamber, allowing the pressures and thus the optical gaps to be adjusted independently. By finding the appropriate pressures for the two chambers one can then line up a transmission peak of each etalon with the desired wavelength. Note that it is not always the transmission peak of the same order that is used.

I will now calculate the amount by which the index of refraction, n , must be changed in order to move a transmission peak by one free spectral range, that is, I will calculate Q_n . Consider on-axis rays for two adjacent orders,

$$m_1\lambda = 2n_1l \quad (2.25)$$

$$m_2\lambda = 2n_2l. \quad (2.26)$$

Subtracting these two equations gives,

$$(m_1 - m_2)\lambda = 2(n_1 - n_2)l. \quad (2.27)$$

But, we assumed adjacent orders, and hence $m_1 - m_2 = 1$. Therefore,

$$n_1 - n_2 = \frac{\lambda}{2l}. \quad (2.28)$$

Hence,

$$Q_n = \frac{\lambda}{2l}. \quad (2.29)$$

The worst case, that is, the circumstances under which the largest change in n will be required, is for the low resolution etalon and the longest observed wavelength. The gap spacing of the low resolution etalon is $l = 0.0201$ cm and the reddest light

that the WHAM spectrometer can study effectively is $\lambda = 7200 \text{ \AA}$. Hence for the worst case, $(\Delta n) = 1.63 \times 10^{-3}$. We thus must change the chamber pressures by enough to achieve this change in the index of refraction.

The increase in index of refraction above the vacuum value ($n_{vac} = 1.0$) for a gas scales approximately linearly with the number density of atoms, and thus for an ideal gas scales approximately linearly with pressure. Hence we can write,

$$n_p = 1 + \beta p, \quad (2.30)$$

where n_p is the optical index of the gas at pressure p measured in atmospheres, and β is the proportionality constant, characteristic of the gas. For air, $\beta \simeq 3 \times 10^{-4}$. The gas we use in our chambers is sulfur hexafluoride (SF_6). The reason for this is that SF_6 has an unusually high value of β , $\beta_{\text{SF}_6} = 7.2 \times 10^{-4}$, and it is also inert and environmentally safe (as opposed to freon, e.g.). Our chambers were designed to withstand 3 atmospheres of pressure, and hence we can change n by 2.2×10^{-3} , safely larger than the value needed for the worst case calculated above.

The relationship between the change in index of refraction of a real gas with the pressure isn't precisely linear as Equation 2.30 would suggest. This is due to attractive inter-molecular forces that cause the pressure to rise more slowly than the number density of atoms. This difference is parameterized using the the Van der Waals' a and b coefficients in the Van der Waals' equation of state:

$$(p + \frac{N^2 a}{V^2})(V - Nb) = NRT, \quad (2.31)$$

where the values of a and b for sulfur hexafluoride are

$$a_{\text{SF}_6} = 0.7857 \text{ Pa m}^6 \text{ mol}^{-2} \quad (2.32)$$

$$b_{\text{SF}_6} = 8.786 \times 10^{-5} \text{ m}^3 \text{ mol}^{-1}, \quad (2.33)$$

and N represents the number of molecules expressed in units of Avogadro's number. The relationship between the pressure and the number of moles of sulfur hexafluoride in a 1 m^3 volume is plotted as the solid curve in Figure 2.8. The dotted line shows the relationship for an ideal gas. In Figure 2.9, the fractional difference between the real (SF_6) and ideal gas cases versus the number of moles N is plotted. One can see from this figure that at the extreme end of the range in N , which corresponds to 3 atmospheres of pressure, there is about a 3% difference. In fact, it is $(\epsilon-1)$ that scales linearly with the number density of molecules, where ϵ is the dielectric permittivity of the gas. The index of refraction, n , is proportional to the square root of ϵ , so we expect the deviation from linearity in n to be about 1.5% at the upper limit of our pressure range.

2.3.4 Scanning versus Imaging

If confined to using a single element detector (commonly a photo-multiplier tube), a natural way to measure a spectrum is to use an aperture to pass only rays close to the axis ($\cos \theta_t \simeq 1$) and then scan either n_g , or l to move the transmission peak through a range of wavelengths and thereby measure the spectrum of the source. The aperture must be small enough so that the $\cos \theta_t$ dependence does not degrade the resolution of the system, and this requirement severely limits the amount of light per time that can be admitted. In scanning l , one moves one of the Fabry-Perot plates with respect to the other while maintaining the very accurate parallelism that is required. This is usually accomplished by using piezo-electric spacers wherein the variation in the length of the spacer is proportional to the applied voltage. One can then scan a range of l by ramping the applied voltage. Alternatively, if the gap is

filled with a gas of some kind, one can then scan n_g by changing the pressure of the gas. This is accomplished by enclosing the entire etalon in a sealed chamber with windows, and then varying the pressure in the chamber. Both of these methods suffer the disadvantage of only accepting light from one spectral element at a time.

The advent of imaging detectors such as charge-coupled devices (CCDs) has allowed a more efficient use of the Fabry-Perot for spectroscopy. Here the variation with θ_t is used to allow simultaneous measurement of multiple spectral elements. This ‘multiplexing advantage’, or more properly ‘multichannel advantage’, leads to a gain in efficiency equal to the number of spectral elements in the spectral range. Another advantage of this method derives from the fact that CCD detectors have much larger quantum efficiencies ($\simeq 4\times$) than photo-multiplier tubes (PMTs), although there can be a disadvantage due to noise associated with reading out the number of electrons in each pixel of the CCD. This method has been demonstrated to give large gains in efficiency in laboratory as well as field tests (see below). In the next section I will describe our ring imaging method in some detail.

2.3.5 Fabry-Perot Spectroscopy via Ring Imaging

The theory and application of ring imaging spectroscopy is discussed in detail by Coakley et al 1996 and Nossal et al 1996. A brief summary is presented here.

Consider the basic condition for constructive interference in a Fabry-Perot system,

$$m\lambda = 2n_g l \cos \theta_t. \quad (2.34)$$

For a given order, m , and for on-axis rays, $\cos \theta_t = 1$, the wavelength $\lambda_o = 2n_g l / m$ is passed by the system. Progressively shorter wavelengths are passed by the system as

we increase the angle of the rays with respect to the optical axis,

$$\lambda_\theta = \lambda_o \cos \theta_t. \quad (2.35)$$

Since there is no dependence on the orientation in a plane perpendicular to the optical axis, a monochromatic light source will produce a symmetric ring pattern if a lens is used to image the angles passed by the Fabry-Perot into positions (see Figure 2.10).

Using Equation 2.23 and Equation 2.35, we can derive

$$\Delta\lambda = (\lambda_o - \lambda) = \lambda_o(1 - \cos \theta_t). \quad (2.36)$$

And thus, converting to a radial velocity scale,

$$-\frac{v}{c} = \frac{\Delta\lambda}{\lambda} = (1 - \cos \theta_t), \quad (2.37)$$

where v is the velocity with respect to the on-axis ($\theta_t = 0$) velocity. As discussed above, we desire our spectrum to cover a range of 200 km s^{-1} . Plugging this into Equation 2.37 and solving for θ_t yields $\theta_t(\text{max}) = 2.09^\circ$. This is a very important number that dictates many elements of the WHAM optical system, as will be seen.

Since the maximum angle passed is $\simeq 2^\circ$, we may safely approximate $\cos \theta_t \simeq 1 - \theta_t^2/2$. Hence,

$$-\frac{v}{c} = \frac{\theta_t^2}{2}. \quad (2.38)$$

We wish to image angles less than $\theta_t(\text{max})$ onto a 1 cm^2 area of the CCD chip. The chip we ultimately acquired for the WHAM system is 2.4 cm on a side, but we didn't know what size chip we would be getting when the optical system was designed. This also allows us to swap in other CCD cameras, already in our possession, which have a 1 cm^2 format, if we need to. Opting for an image that is 2.4 cm on a side also generates

problems caused by the need for very large binning factors (and the resulting increase in noise) as well as problems related to the fact that the generated electrons would be spread over a larger area and the charge transfer efficiency is lower for lower charge levels. For our 1 cm² ring image, we can calculate the effective focal length of the ring imaging lenses, f_{eff} ,

$$0.5 \text{ cm} = f_{eff} \tan \theta_t(max) \quad (2.39)$$

$$f_{eff} = 13.7 \text{ cm}. \quad (2.40)$$

Note that the aperture defined by the etalons has a diameter of 15cm (the largest available size) and thus this f_{eff} implies an f-number $\simeq 0.9$. The actual imaging of the rings is done with more than one lens, and it was a non-trivial problem to design the optical system to meet the imaging specifications required for this application (see below).

I will now derive the relationship between velocity and position on the CCD chip. Let r be the polar radius vector measured from the center of the ring pattern on the chip (see Figure 2.10). Then,

$$r = f_{eff} \tan \theta_t \simeq f_{eff} \theta_t \quad (2.41)$$

implies

$$\frac{v}{c} = \frac{r^2}{2f^2}. \quad (2.42)$$

One important consequence of this relationship is that equal spectral intervals, $\delta\lambda$, cover equal areas on the CCD chip. Let dA represent an infinitesimal element of annular area

$$dA = 2\pi r \, dr. \quad (2.43)$$

Since,

$$r \, dr = \frac{f_{eff} \, dv}{c}, \quad (2.44)$$

Then,

$$dA = \frac{2\pi f_{eff}^2}{c} \, dv. \quad (2.45)$$

Equation 2.45 is independent of r , and thus the annuli corresponding to equal spectral intervals get progressively narrower as one moves away from the center of the pattern in such a way as to make the annular area constant. This leads to the peculiar optical specification that the tolerance for image aberrations is significantly more stringent at the edge of the field than at the middle. The area on the chip of an annulus corresponding to an 8 km s^{-1} velocity interval is, $dA = 3.15 \times 10^{-2} \text{ cm}^2$. At the edge of the field, this must equal $2\pi r \, dr$ with $r = 0.5 \text{ cm}$, which leads to a spatial width for the outermost spectral element $(dr)_{edge} = 100 \text{ } \mu\text{m}$. This dimension defines the required image quality and determines the degree to which on-chip binning of the CCD can be employed (see next section).

2.3.6 CCD Detection of the Ring Pattern

To measure a spectrum, we use the chamber pressure to adjust the optical gaps, such that the red edge of the desired wavelength range is passed for rays on the optical axis. An exposure is then started and electrons accumulate in the CCD pixels from the absorbed photons which are distributed radially across the chip according to their wavelength. After the exposure is finished, the number of electrons in each pixel is read out and stored on the computer disk. The spectrum is then reconstructed in software by averaging the counts in each annular bin in order to form a one-dimensional plot of intensity versus wavelength.

There is a limit to the accuracy with which the number of electrons in a pixel can be determined. The error in the determination of the number of electrons is termed the *read noise* of the CCD system and is a fixed number, independent of the number of electrons in the pixel. For this reason, it would be advantageous to reduce the number of pixels that need to be read within each equal area annulus. This is possible through on-chip binning of the pixels. It is possible to combine pixels on the CCD chip to form super-pixels of arbitrary size. Then the total number of electrons falling within a super-pixel are read as one with nearly the same noise as in the reading of a single pixel (we have discovered that the read noise is slightly higher for binned pixels, an effect which I discuss in the next chapter). The optimal super-pixel size is the largest size that will not significantly degrade the ring imaging resolution. This means that the super-pixel cannot be larger than the width of the outermost ring, which was calculated above to be $(dr)_{max} = 100 \mu m$. The size of a pixel on our CCD chip is $24 \mu m$ and hence 4×4 binning is optimal. This leads to $16 \times$ fewer reads and thus, in a regime where the read noise is dominant, a factor of four reduction in the noise. We actually do suffer a decrease in spectral resolution near the edge of the chip due to the fact that the pixel width is a non-negligible fraction of the ring width, but this liability is more than outweighed by the reduction in read noise afforded by using this large super-pixel.

2.4 The WHAM Optical design

The requirements of the WHAM optical system are to accept light from a 1° beam on the sky, use this light to illuminate the 15 cm etalons, with light filling angles corresponding to a 200 km s^{-1} velocity interval, and then to image the resulting ring

pattern onto a 1 cm^2 area of the CCD chip. We also need to pass the light through a narrow-band interference filter along the way in order to eliminate unwanted orders that are passed by the two Fabry-Perots. Further, we need to have the option of imaging the sky onto the CCD chip for verifying our look direction and also to allow the possibility for velocity-resolved imaging work at $\sim 1'$ resolution. This section describes the optical design we have developed to accomplish these goals. Note that the conceptual design for the optical system was done by Ron Reynolds and the detailed optical design was worked out by John Harlander of St. Cloud University using the Code V ray tracing program.

2.4.1 Overview of the Optical System

A schematic drawing of the overall WHAM system hardware is shown in Figure 2.11. The two main parts of the system are the siderostat and the WHAM spectrometer. On the left is depicted the all-sky siderostat which contains two folding mirrors and a 0.6 m diameter, 8.6 m focal length lens. On the right is the WHAM spectrometer, which along with the computer systems and various control systems is housed in a small trailer.

Light from a chosen place on the sky enters the siderostat and is guided by the siderostat folding mirrors into the spectrometer. After passing through the 0.6 m telescope lens the light travels along a tube into the WHAM trailer. The optical path crosses the trailer near the ceiling, then reflects off a folding mirror that sends it down through the two Fabry-Perot etalons chambers. Next, the light passes through the ring imaging lens just after the etalons, reflects off another folding mirror, and an image of the ring pattern is then formed at the location of the interference filter. From

there, the ring pattern is reimaged by a sophisticated 4 element camera lens onto the CCD chip, where the image is recorded. When an image of the sky is desired, three additional lenses are inserted between the last folding mirror and the filter which have the effect of imaging the sky onto the interference filter. Since the interference filter plane is imaged onto the CCD chip, this images the sky onto the detector.

2.4.2 The Siderostat

The purpose of the WHAM siderostat is to steer light from an arbitrary point in the observable sky along the optical axis of the spectrometer. The siderostat consists of two 90° folding mirrors that are mounted on perpendicular rotatable axes. The siderostat structure also houses the 0.6 m telescope lens, located between the last mirror and the light tube.

The two movable axes form an altitude-azimuth mounting system, but tipped over such that the zenith of this “alt-az” system is pointing toward the northern horizon. Our tipped system needs only 180° rotations for both axes to cover the entire sky, avoids the need for more than 360° rotations of a conventional alt-az, and doesn’t have a zenith exclusion zone problem. The problem still exists, but for us is located at the northern and southern horizon points. It is never advisable to observe near the northern horizon point since an object there is always better observable at another time, and having a tracking problem at the southern horizon point is clearly preferable to having one at the zenith.

The siderostat hardware was developed at Physical Sciences Laboratory and the control software was developed as a joint project between PSL and the WHAM team.

2.4.3 The Pre-etalon Optics

The desire for a 200 km s^{-1} spectral interval requires a 2.09° half angle for the cone of rays passed through the etalons (Section 2.3.5). Our desire for a 1° beam on the sky requires that we map the 2.09° half angle cone onto a 0.5° half angle cone. This is the purpose of the pre-etalon optics.

The basic idea is to add an additional lens of appropriate size and focal length in front of the etalons (see Figure 2.12). This ‘telescope’ lens images the sky onto a plane between the etalons, which has the added benefit of separating the spatial and spectral information. In many Fabry-Perot systems there are no pre-etalon optics and thus the detector simultaneously records the Fabry-Perot ring pattern and an image of the sky. Each annular ring on the detector corresponds to a particular spectral interval, as before, but in these systems the amount of light reaching that ring on the detector is now determined by the corresponding spatial ring on the sky. Hence, an ambiguity exists, and a bright spot on the detector may indicate either a spectral feature, a spatial feature, or some combination of the two. With the WHAM system, the light from each point on the sky is spread uniformly over the CCD detector, and no such ambiguity exists.

The conservation of etendue (area $\times$ solid angle product through the optical system) requires the telescope lens to be $15 \text{ cm} \times 2.09 / 0.5 \simeq 63 \text{ cm}$ in diameter. In order to illuminate the full area of the etalons, the focal length of the telescope lens must satisfy $\tan 0.5^\circ = 15 \text{ cm} / f_{L1}$. This implies $f_{L1} \simeq 860 \text{ cm}$. Without the use of an additional lens in front of the etalons, the cone of rays illuminating outer parts of the etalons would be tilted on average with respect to the optical axis. Since only light

within 2.09° of the optical axis is passed by the rest of the optical system, the net effect is to selectively underilluminate the blue side of the spectrum. For this reason, we add an additional lens that is placed just ahead of the etalons with nearly the same focal length as the telescope lens which results in parallel (telecentric) illumination of the etalons.

2.4.4 The Post-etalon Optics

The Fabry-Perot etalons pass different wavelengths of light at different angles. Hence after passing through the etalons, the spectrum of the light is encoded in angles according to Equation 2.35. The primary purpose of the post-etalon optics is to map these angles into positions so that the spectrum can be recorded. The simplest method is to place a lens just after the etalons and then place an imaging detector in the focal plane of the lens. However, we also need a narrow-band interference filter in our system to eliminate unwanted orders passed by the Fabry-Perot. Because of the unacceptably high cost of good quality 15 cm diameter filters, we have opted to image the ring pattern at an intermediate location in order to allow the use of 7.5 cm interference filters. Our post-etalon optical design is schematically illustrated in Figure 2.13.

The lens just after the etalons, L3, serves to image the ring pattern onto the filter. The 7.5 cm filter size along with the 2.09° maximum angle determine the focal length, $\tan 2.09^\circ = 1.5 \times 2.54 \text{ cm} / f_{L3}$ implies $f_{L3} \simeq 104 \text{ cm}$. This lens in combination with L2 also serves as a field lens, imaging the entrance pupil (L1) onto the interference filter aperture. The next lens in the optical path is L4 which serves the purpose of parallelizing the ray cones for passage through the interference filter. The lens L5 along with L4 also acts as a field lens to image the etalons, and hence the sky, onto

the final set of lenses L6 - L9, which we collectively refer to as the camera lens. The four element camera lens focuses the ring image onto a 1 cm^2 region of the CCD chip, where the spectrum is recorded. The conservation of etendue requires that the cone of rays incident upon a particular point on the chip has a half-angle of $2.09 \times 15 \text{ cm} / 1 \text{ cm} \simeq 30^\circ$. Hence the camera lens must have an $f/\# \simeq 0.9$. Due to chromatic aberration in the system, the camera lens is mounted on a movable track, and its precise distance from the chip is adjustable with a micrometer dial, whose setting is different for each emission line. For example, 1.6 mm of travel is required to move between $\text{H}\alpha$ ($6563\text{\AA}$) and $\text{H}\beta$ ($4861\text{\AA}$) (see Section 3.9). The lens diagrams shown in Figures 2.12 and 2.13 are schematic. In order to demonstrate the true complexity of the ring imaging system, I have included Figure 2.14, which shows a scale drawing of the actual camera lens system. Three sets of rays are traced through the lenses to the CCD chip, corresponding to on-axis rays, rays tilted by 1.48° off-axis at the etalons, and rays tilted by 2.09° off-axis at the etalons (the blue edge of our bandpass).

2.4.5 The Sky Imaging Lenses

In order to image the sky on the CCD chip instead of the spectral ring pattern, lenses can be inserted into the beam between L3 and L4. Three lenses are used, and they are mounted on a motor driven movable table that slides into and out of the beam. These lenses take the sky image that is between the etalons and reimage it onto the filter IF, which as before is then imaged onto the CCD. An additional feature of our arrangement of these lenses is that the spectrum is imaged on an intermediate plane where we have placed an adjustable iris. Hence, when we are imaging the sky, we can adjust the wavelength interval that is passed from a minimum value equal to

the instrumental resolution of 12 km s^{-1} up to a maximum value equal to the full spectral range, which is 200 km s^{-1} .

References

Coakley, M. M., Roesler, F. L., Reynolds, R. J., & Nossal, S. 1996, *Applied Optics*, Vol. 35, No. 33, 6479

Nossal, S., Roesler, F. L., Coakley, M. M., & Reynolds, R. J. 1997, *J. Geophys. Res.*, in press

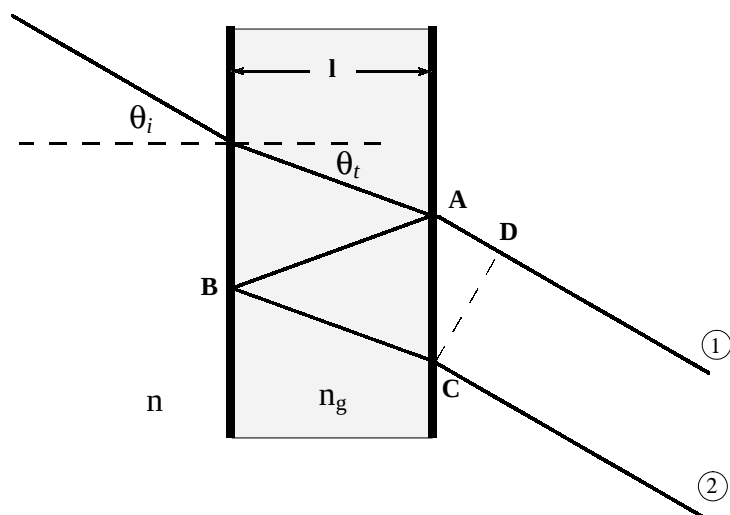


Fig. 2.1.— The optical path difference between consecutive rays.

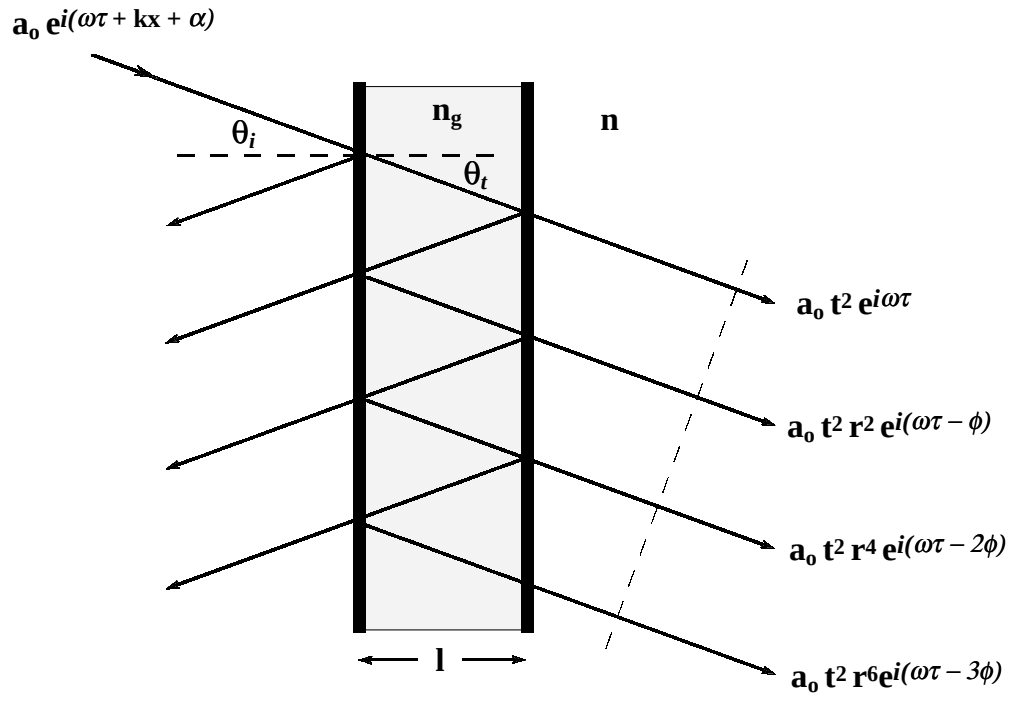


Fig. 2.2.— A Simple Fabry-Perot with multiple reflections between the plates.

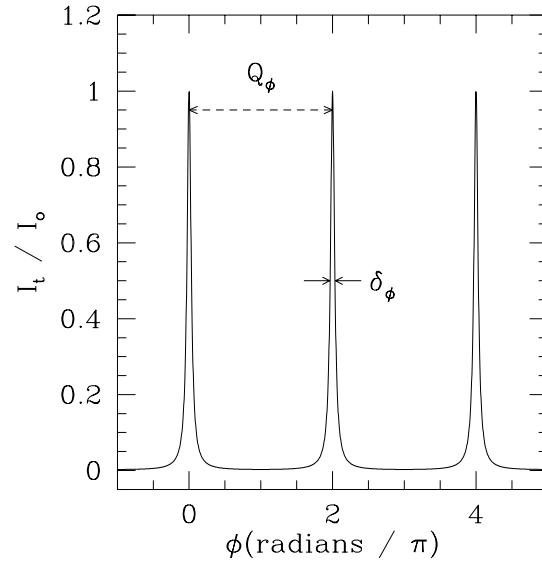


Fig. 2.3.— I_t / I_o versus ϕ for an ideal Fabry-Perot with $R = 0.9$.

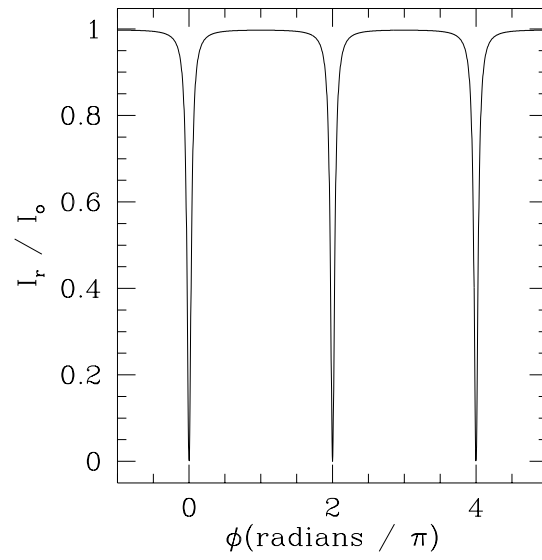


Fig. 2.4.— I_r / I_o versus ϕ for an ideal Fabry-Perot with $R = 0.9$.

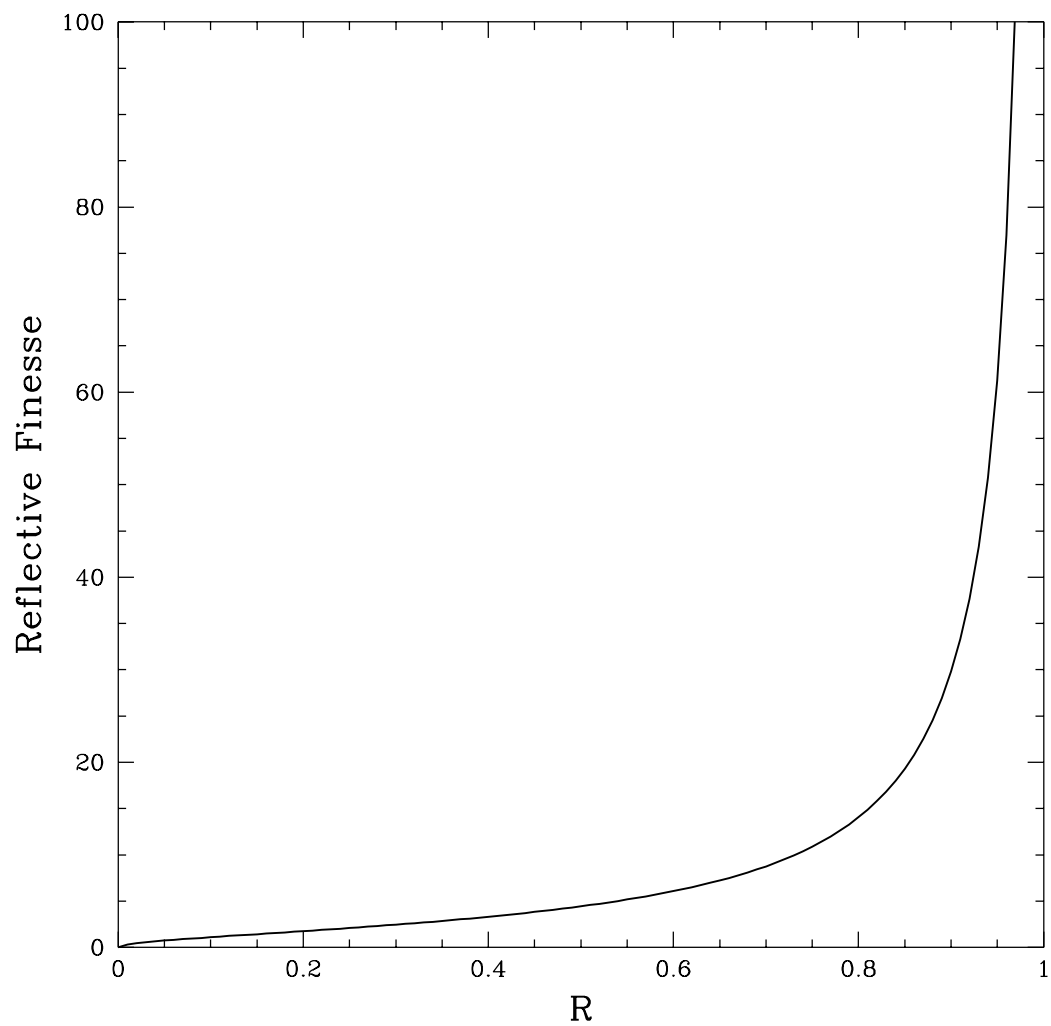


Fig. 2.5.— Reflective finesse vs. R .

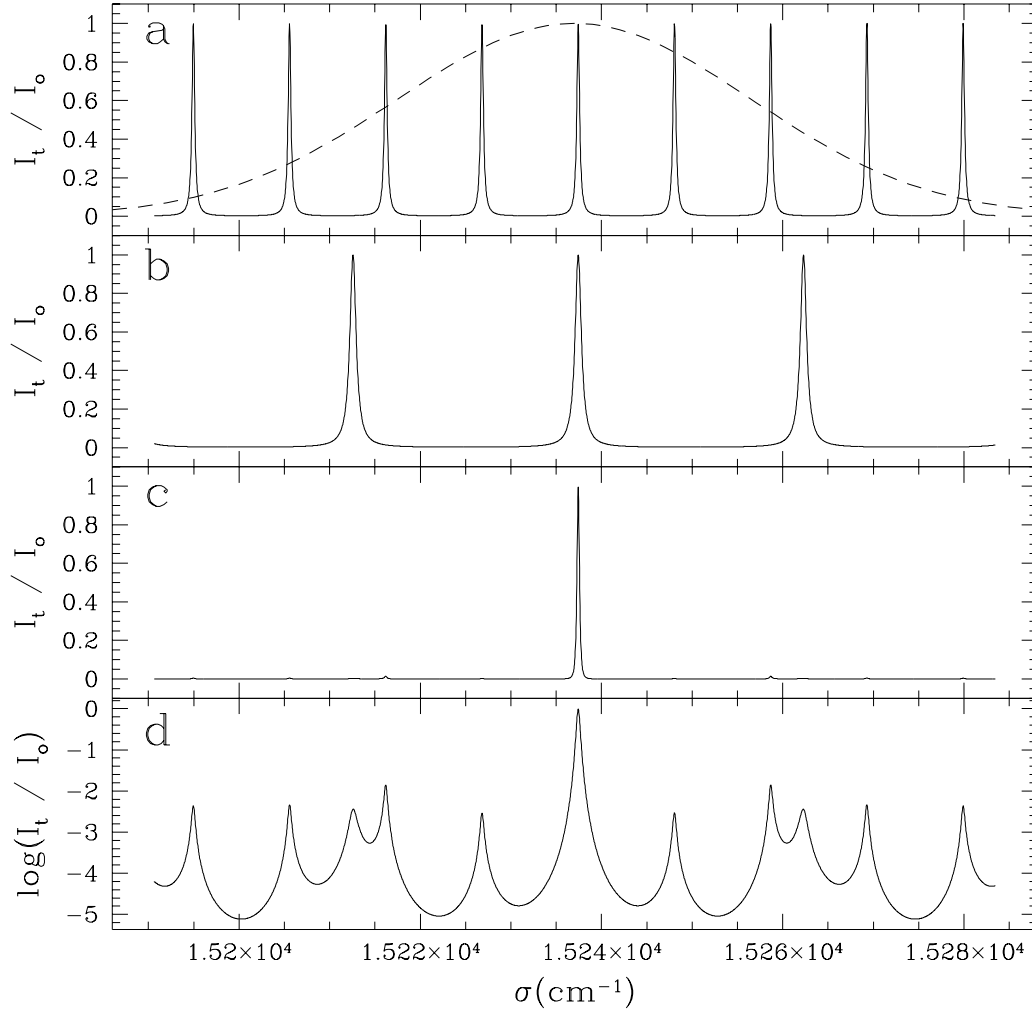


Fig. 2.6.— The theoretical transmission functions of the WHAM etalons: a) solid line: the high-resolution etalon, dashed line: interference filter, b) the low-resolution etalon, c) the product of the etalon transmission functions, d) same as c) but on a log scale.

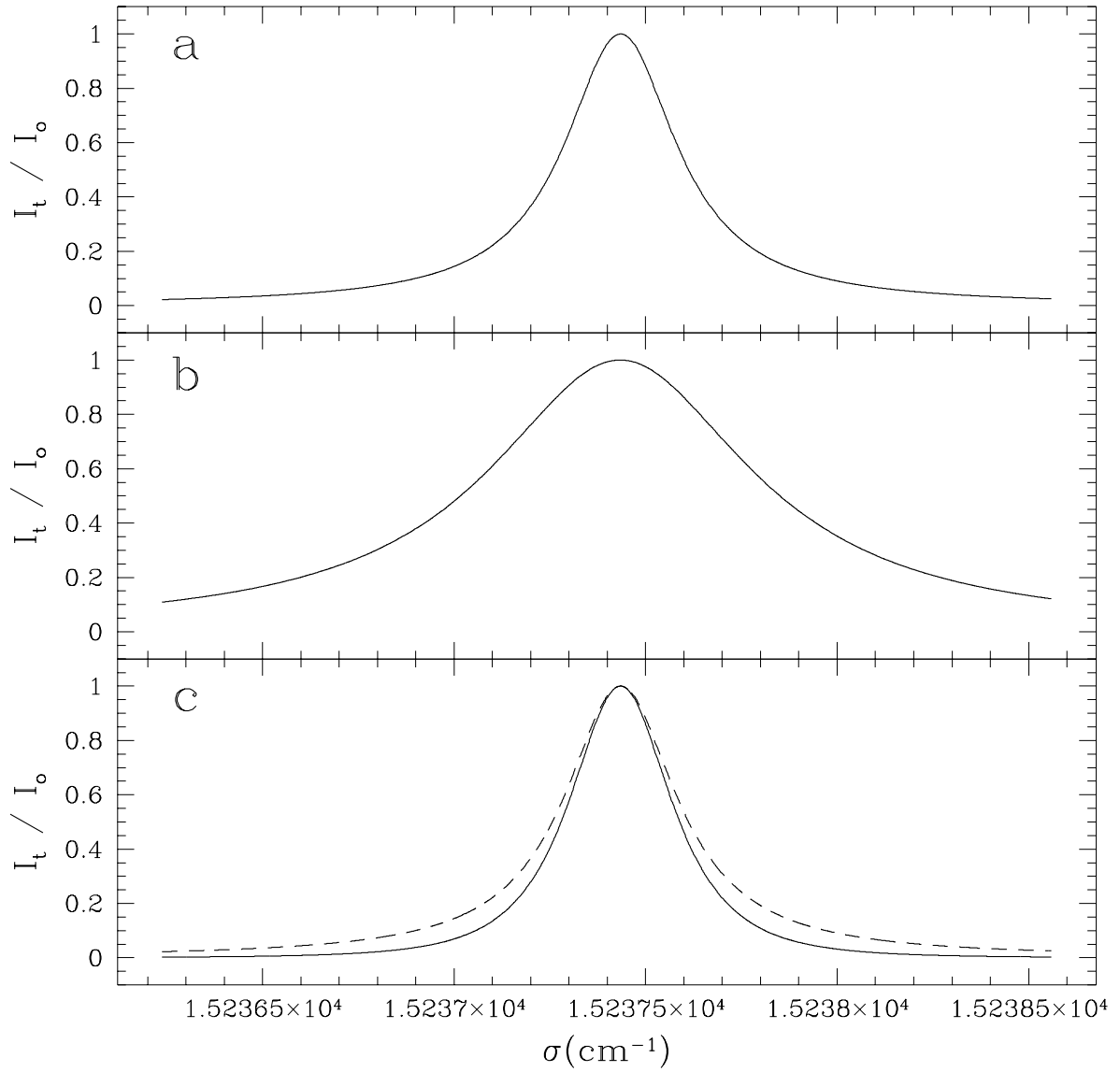


Fig. 2.7.— The theoretical transmission functions of the WHAM etalons: a) the high-resolution etalon, b) the low-resolution etalon, c) solid line: the product of the two, dashed: the high-resolution etalon alone.

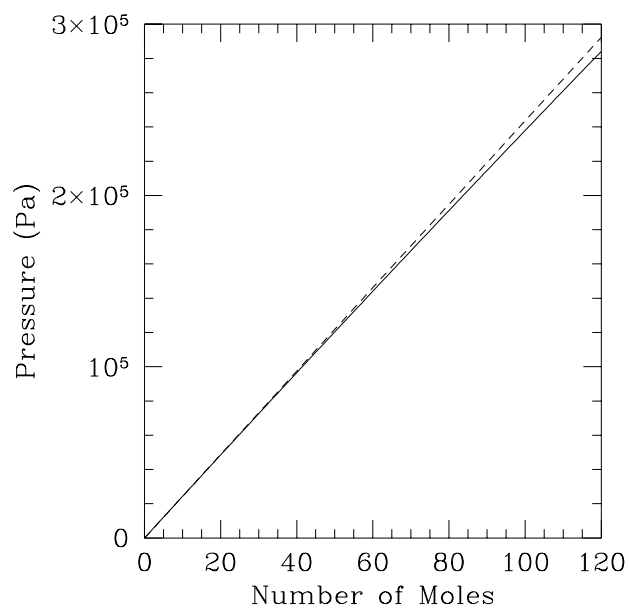


Fig. 2.8.— Sulfur hexafluoride compared to ideal gas, solid line: pressure vs number of moles of SF₆ gas in a 1 m³ volume, dotted line: same for an ideal gas.

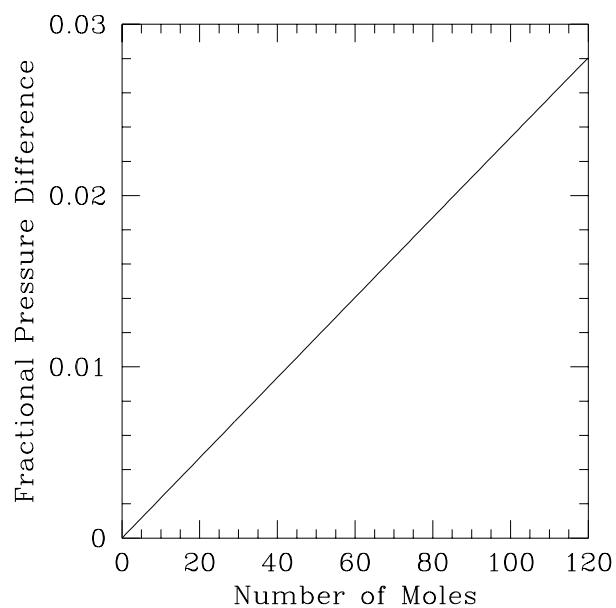


Fig. 2.9.— Sulfur hexafluoride compared to ideal gas, fractional difference between SF₆ and an ideal gas vs. number of moles in a 1 m³ volume.

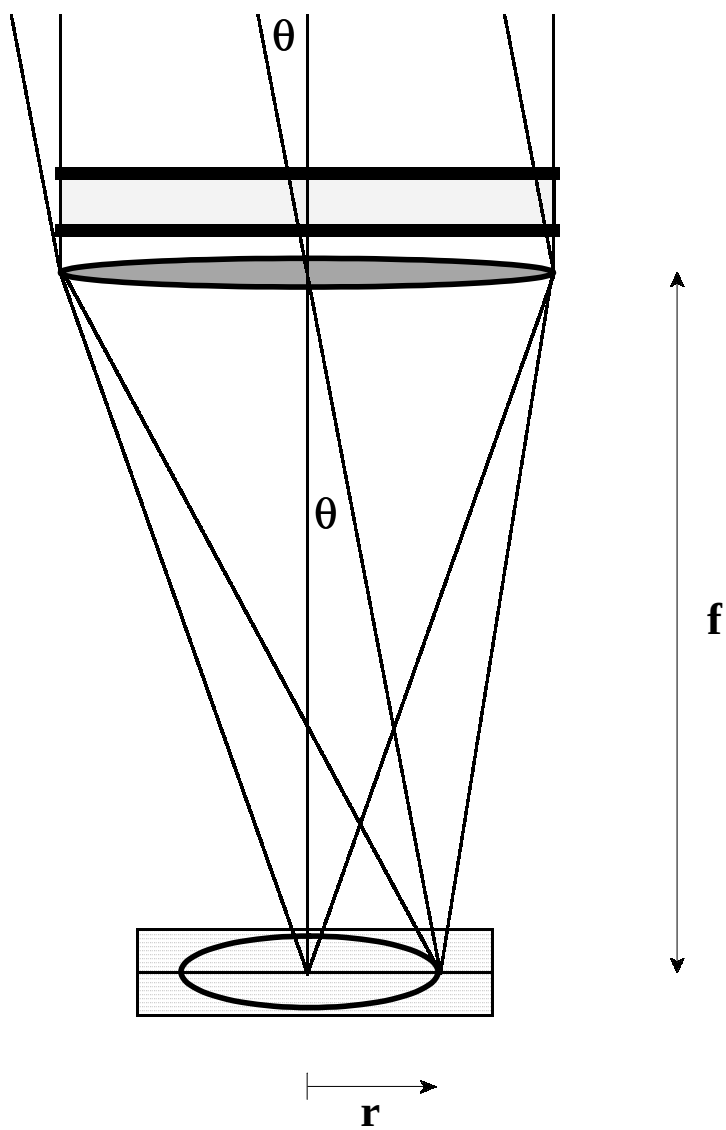


Fig. 2.10.— Fabry-Perot ring imaging.

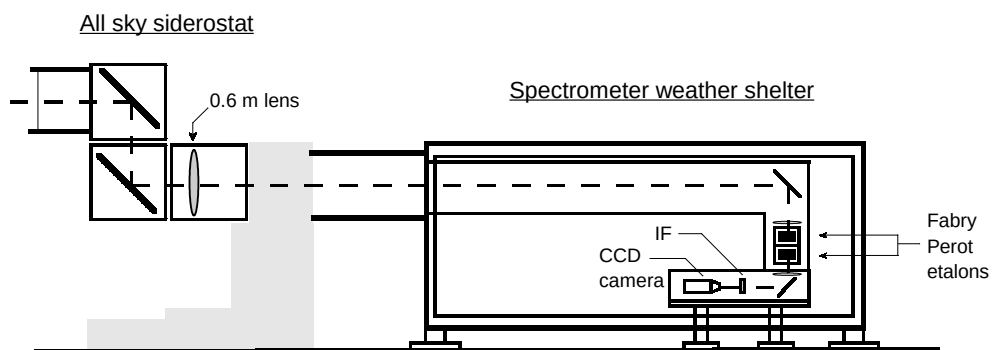


Fig. 2.11.— The WHAM system.

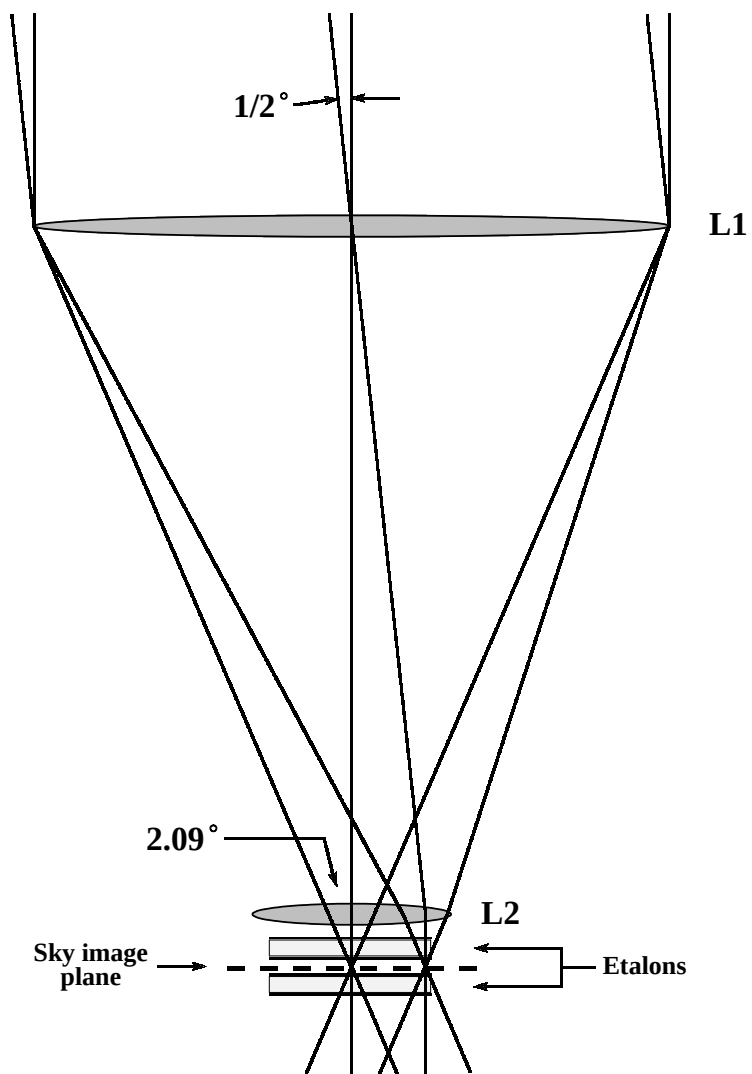


Fig. 2.12.— The WHAM pre-etalon optics (schematic).

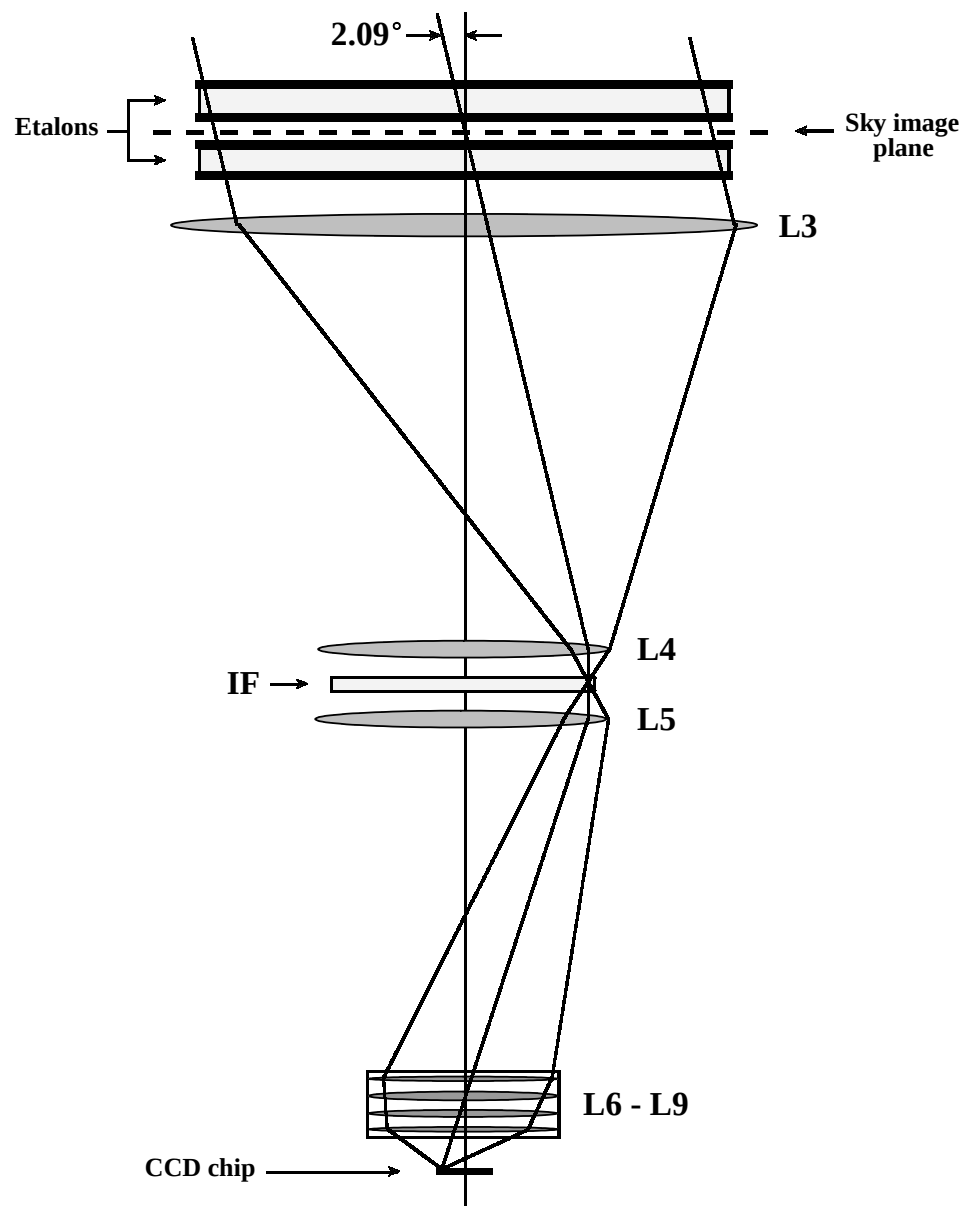


Fig. 2.13.— The WHAM post-etalon optics (schematic).

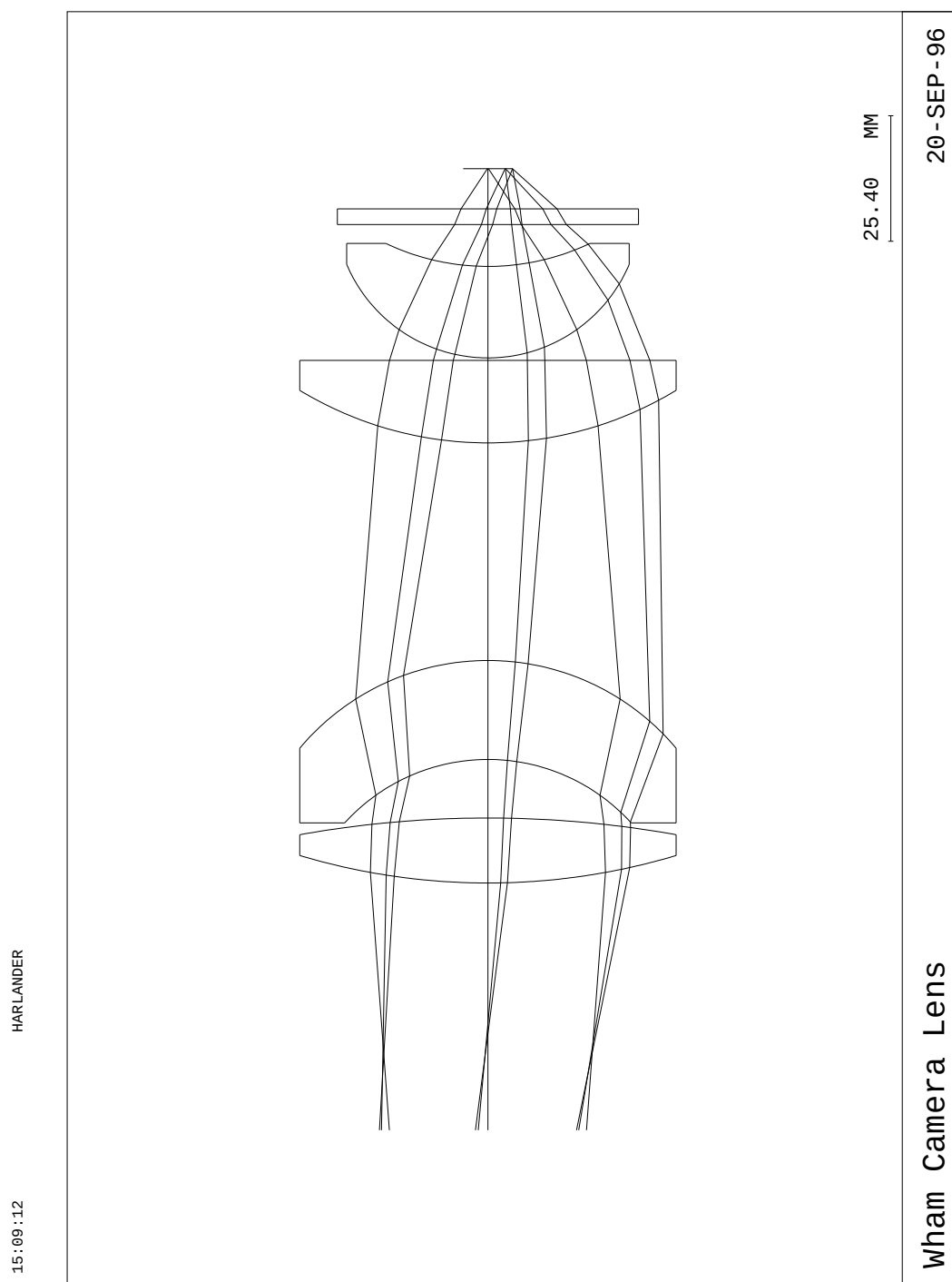


Fig. 2.14.— The WHAM camera lens.

Chapter 3

The WHAM Spectrometer: Performance Characteristics

3.1 The Instrument Profile

The instrument profile was discussed from a theoretical perspective in Chapter 2. Here I present measurements of the WHAM instrument profile and discuss some of the important factors that affect its width. By definition, the instrument profile is the response of the instrument to a delta function input spectrum. In practice, we measure the instrument profile by measuring the spectrum of a narrow line, used as an approximation of a delta function. So long as the line width is a small fraction of the instrumental width, this procedure will yield accurate results.

The narrow line used for this study is a bright line near $\lambda 6300$ emitted by a simple neon chassis light which is passed by our [O I] $\lambda 6300$ interference filter. The line width of this source was measured to be $\simeq 1.4 \text{ km s}^{-1}$ (Roesler, private communication). The chassis light was placed above the etalons and passed through two separated

diffusers in order to form a source that was uniform in both space and angle. The data points in Figure 3.1 show the measured spectrum of this neon line. The curve shows the theoretically expected shape which is the product of two Airy functions with different resolving powers (see Section 2.3.2). The curve was adjusted to match the central part of the line and is not a fit. Overall, the shapes match quite well, but there is a slight asymmetry on the blue wing of the measured profile due to etalon defects. The only adjustable parameter in the Airy function ‘model’ is the reflectivity. The reflectivity required to match the width of the measured profile is $R = 0.85$. In fact $R \simeq 0.90$; a lower R is necessary to fit the data because the instrument profile has been broadened by effects not included in the simple Airy function model. Properties of the spectrometer system which will tend to broaden the instrument profile include non-flatness of the etalon plates, nonparallelness of the etalon surfaces, slight misalignments of the optical axes of the two etalons, pixel sampling effects, variations in the reflectivity of the etalon surfaces, and the “point spread function” of the ring imaging system. The overall broadening by the sum of these effects is still not large. The line shown in Figure 3.1 has a line width of 9.3 km s^{-1} FWHM, only slightly wider than the $\simeq 7.0 \text{ km s}^{-1}$ FWHM width expected for an ideal Fabry-Perot system with our gap spacings and reflectivity. However, note that this spectrum was taken with 1×1 binning and with the etalons masked to $\simeq 3.5 \text{ cm}$ aperture and that the instrument profile in our normal operating conditions is somewhat broader (see below).

I carried out investigations of a few of the important instrument profile broadening mechanisms. Working with Nikki Hausen, an undergraduate on the WHAM project, I gathered a collection of ring images of the neon line under different circum-

stances. Exposures were taken at a range of chamber pressures at constant tune (i.e. pressure difference) to look for a radial dependence of the instrument profile. We also collected exposures for 1×1 binning and 4×4 binning, both with the etalons unmasked and with them masked down to sample a smaller region of the etalons. The masking was done by placing an opaque screen containing an approximately 3.5 cm circular hole just above the etalons. The images were all processed into spectra after carefully finding the ring centers, and then the instrument profile width was measured for each spectrum.

Figure 3.2 shows a plot of the widths (FWHM) of the instrument profiles versus ring position, determined for the case of 1×1 binning with the etalons masked. The widths are fairly constant at $\simeq 9 \text{ km s}^{-1}$ but may show a slight downward trend. Compare this with the widths shown in Figure 3.3 measured for the case of 1×1 binning but with the etalons unmasked, i.e. using the full aperture of the etalons. The widths in the latter case are slightly over 10 km s^{-1} , indicating an $\simeq 1 \text{ km s}^{-1}$ broadening caused by non-uniformities in the etalon plates.

Figure 3.4 shows the widths determined for the case of 4×4 binning with the etalons masked. Here we see a clear upward trend to the data points as the ring moves toward the edge of the chip. This effect is caused by the fact that near the edge the pixel size becomes comparable to the ring width (in fact, at the very edge they are equal). In the process of forming a spectrum from a ring image, the chip area is divided into a series of concentric annuli of equal area (see Section 2.3.5). Pixels are chosen for a given annulus based on whether the center of the pixel falls within the annulus; portions of the area of the selected pixels will be outside the desired annular area. This pixel sampling effect is an additional source of broadening that

must be convolved with the previously calculated Airy function to obtain the expected instrumental profile. Near the middle of the ring pattern this effect is not important since the annuli corresponding to a spectral resolution element are much wider than the pixel size. Thus, the severity of this broadening mechanism increases with radius from the center of the ring pattern, as is seen in the data. One aspect of this study that I do not understand is that the widths of the 4×4 binned images shown in Figure 3.4 are $\simeq 0.6 \text{ km s}^{-1}$ wider than for the 1×1 images shown in Figure 3.2 near the center of the chip, where the 4×4 pixel sampling effect should be small. Figure 3.5 shows the width of the images taken using 4×4 binning and with the etalons unmasked. This is the normal configuration used when collecting science data. In this case, we have both the $\simeq 1 \text{ km s}^{-1}$ broadening caused by non-uniformities in the etalon plates and the pixel sampling effect near the edge of the ring image, resulting in instrument profile widths of $\simeq 11\text{--}12 \text{ km s}^{-1}$.

3.2 Wavelength Calibration

The Fabry-Perot etalons pass different wavelengths of light at different angles according to Equation 2.35. As explained in Section 2.3.5, in a perfect imaging system this equation leads to the rule that equal area annuli on the CCD chip correspond to equal spectral intervals. The purpose of this wavelength calibration study is to determine the wavelength interval that corresponds to a given annular area on the chip. This study will also check the equal area annuli rule and thus check the performance of the ring imaging system in linearly mapping the angle of a ray at the etalons to a distance from the ring pattern center on the chip. This study does not provide an absolute wavelength calibration (i.e., the zero of the velocity scale), it only determines

wavelength intervals as a function of radius on the chip. For the $H\alpha$ line, an absolute calibration is available in the form of the geocoronal $H\alpha$ emission line which appears in every spectrum and provides the zero of the velocity scale. In the case of the [S II] $\lambda 6716$ line, there is a nearby neon line that can be used to set the zero point. For the [O I] $\lambda 6300$ line, there is an atmospheric air glow line that can be used. In some cases, a laboratory lamp must be used to set the zero of the scale.

The method employed in this study is to measure the spectrum of a source containing two spectral lines of known wavelength separation and then to use this spectrum to set the velocity scale. The source we used is a hollow cathode lamp containing both hydrogen atoms and deuterium atoms. The Balmer α line of hydrogen ($H\alpha$) differs in wavelength from the Balmer α line of deuterium ($D\alpha$) by $1.78 \text{ \AA}$ which corresponds to a velocity difference of 81.55 km s^{-1} . This is a convenient wavelength difference since the spectral bandpass of the WHAM ring image is 200 km s^{-1} .

In order to capitalize on this convenient pair of lines, we had a special hollow cathode lamp made which contains both hydrogen and deuterium. With an arrangement similar to that used for the instrument profile observation, I obtained images at a range of pressure values to produce the spectra shown in Figure 3.6. The square data points highlight one of the spectra. The tall peak on the left is the $D\alpha$ line, and the shorter peak on the right is the $H\alpha$ line. Each data point in this spectrum is the result of an average over 85 pixels on the CCD chip. The right most data point in this spectrum is the average of a disk of pixels at the location of the ring center. The next data point is derived from an annular region containing 85 pixels surrounding this disk, and so on. An approximate wavelength scale, which assumes a linear relationship between data point number and wavelength determined in a previous calibration

effort, is shown on the abscissa. The zero of this velocity scale is arbitrary. Also shown in this figure are the spectra derived from images taken at a range of pressures (increasing the pressure in the etalons while holding the pressure difference between the etalons constant causes the rings to expand). Note that the right most spectrum is partly clipped by the edge of the bandpass and that a few of the left most spectra are falling off the edge of the aperture.

Each of the spectra shown in Figure 3.6 was fitted with Gaussian components and a parabolic background. Two Gaussians were allotted to each peak in the spectra, a wide one and a narrow one, and the centroid of the narrow one was used to locate the line in the subsequent analysis. This prevents subtle changes in the shapes of the line wings from distorting the positions measured for the peaks. Figure 3.7 shows the measured position of the $D\alpha$ peak versus the pressure in the chamber containing the high-resolution etalon, and Figure 3.8 shows the same plot for the $H\alpha$ peak. The points at higher pressures correspond to the spectra nearer to the left in Figure 3.6 and are derived from rings further from the center of the aperture. Both graphs show a linear relationship between wavelength and pressure and both have slopes corresponding to $\simeq -2.8 \text{ km s}^{-1}$ change in velocity per 1 cm change in pressure.

Figure 3.9 shows the difference between the $H\alpha$ peak position and the $D\alpha$ peak position versus pressure. Notice that the vertical scale is greatly expanded, and thus this graph is showing an approximately constant value of $\simeq 80 \text{ km s}^{-1}$ for the difference. Very interestingly, this graph does show, however, that there is a systematic distortion of the wavelength scale. Except for the points at the extremes of pressure which correspond to spectra not well fit due to the fact that they are partly off the aperture, the data points show a clear, approximately linear trend. The difference between the

two measurements for the same pressure at $p = 79$ cm is an indication of the error on these data points. Figure 3.10 shows a linear fit to the data points between a pressure of 70 cm and 100 cm. The slope of this fit line is -0.048 km s^{-1} per cm. Since the real difference in wavelength between the two lines is a constant, the downward slope of the data points indicates that the ring imaging optics are compressing the outer parts of the image with respect to the inner parts. This distortion can easily be removed in software by including the effect in the formula mapping data point number to velocity, but at present it isn't clear that the linear trend shown here extends over the whole spectral range. The $\text{H}\alpha$ - $\text{D}\alpha$ wavelength separation is too wide to map out this distortion over the full bandpass; I was only able to sample an interval of $\simeq 85 \text{ km s}^{-1}$ near the center of the bandpass. A future study will use two thorium lines with a closer spacing in wavelength to map out this distortion across a larger fraction of the 200 km s^{-1} spectral window.

3.3 Imaging Performance

The WHAM spectrometer is capable of two types of imaging, ring imaging and sky imaging. Ring imaging refers to the mapping of the angle encoded spectral information at the etalons into rings on the CCD. Since WHAM is primarily a spectral instrument, this type of imaging is the more important of the two. Sky imaging refers to the capability of imaging the sky on the CCD chip by inserting extra lenses into the beam. This is primarily for verifying pointing for spectra but also provides the capability of gathering higher spatial resolution ($\sim 3'$), velocity resolved, narrow-band images of the ionized gas in the diffuse interstellar medium.

3.3.1 Ring Imaging Performance

In order to test the performance of the ring imaging system, we used a single very high resolution Fabry-Perot etalon together with a thorium hollow cathode lamp. The etalon used had a surface reflectivity $R = 0.90$ and a gap spacing of 3.3 mm. This leads to a resolving power $\mathcal{R} \simeq 300,000$ (see Equation 2.16). Increasing the gap spacing of the etalon decreases the free spectral range. Hence, with this larger gap etalon, many free spectral ranges fall within our aperture. The finesse, however, depends only on the reflectivity and thus is the same as for our regular etalons, implying that the instrument profile width is narrower by the ratio of the resolving powers. Our usual instrument profile width is 8 km s^{-1} , and thus with the wide gap etalon, the instrument profile is $\simeq 1 \text{ km s}^{-1}$ wide. An 8 km s^{-1} line maps to a ring that is $100 \mu\text{m}$ wide at the edge of the field for our regular etalons, and hence with the very high resolution etalon we expect an infinitely narrow line to produce rings that are about $12 \mu\text{m}$ wide at the edge, which is about one-half the side of a 1×1 pixel. Hence, the very narrow line produced by the thorium lamp will produce a ring on the CCD chip that is spatially narrow enough to test the imaging performance.

Figure 3.11 shows a portion of an image of rings produced by illuminating the very high resolution etalon with a thorium line near $\text{H}\alpha$ from a hollow cathode lamp. The CCD camera was operated at 1×1 binning for this exposure to maximize the spatial resolution with which the ring pattern is measured. Notice that there are approximately nine free spectral ranges within our bandpass. Within each free spectral range we measure a bright line and a fainter line on the blue side of the bright line. A closeup of the upper right portion of the ring pattern is shown in Figure 3.12. It is

apparent from this image that these rings are approximately 2 pixels wide.

A horizontal slice near the center of the thorium ring image was formed (averaged over twenty rows to decrease the noise) and is shown in Figure 3.13. The plot depicted here shows only a small portion of the slice, taken from near the extreme right side. One can see the spikes from the bright line along with the accompanying fainter line peaks. By fitting these peaks I have determined that the peaks are between 1.5 and 2 pixels wide. Since each 1×1 pixel is $24 \mu\text{m}$ across, this corresponds to a distance of $\simeq 40 \mu\text{m}$.

A sophisticated ray tracing program called Code V was used by John Harlander of St. Cloud State University to optimize the WHAM lens design. This program is also capable of predicting the imaging performance of the optical system. Figure 3.14 shows a diagram indicating the expected ring imaging performance of the WHAM optical system at various off-axis angles. The Code V program generates a uniform grid of rays at the etalon position, all at a specified angle with respect to the optical axis of the etalon. The program then traces the rays through the lens system to the CCD chip. In an ideal optical system, all of these rays would map to a single point on the detector. The top splotch in Figure 3.14 shows the final locations of the grid of rays tilted at 2.09° with respect to the optical axis of the etalons (the blue edge of the spectrum). As our system is not perfectly achromatic (see Section 3.9), the imaging performance depends somewhat on the wavelength of the light, and the diagrams in the figure are appropriate near $\text{H}\alpha$. The RMS scatter of these points is $34 \mu\text{m}$, but it is obvious that the scatter is larger in the horizontal direction, which corresponds to the azimuthal direction on the ring image, and smaller in the vertical direction, which corresponds to the radial direction. This asymmetry was designed

into the system since radial spreading broadens the instrument profile much more than spreading tangent to the ring. The full spread of these points in the radial direction is $\simeq 30 \mu\text{m}$. This is smaller than but comparable to the $40 \mu\text{m}$ measured value of the ring width. The difference is probably due to slight misalignments in the real optical system, as well as slight deviations from ideal in the positioning of the lenses along the optical axis. An 8 km s^{-1} spectral interval corresponds to a ring width of $\simeq 100 \mu\text{m}$ at the edge of the chip, indicating that we expect only a slight broadening due to the imperfect imaging at the edge of the ring image field.

The splotch in the middle of Figure 3.14 shows the spot diagram resulting from rays tilted at 1.48° off-axis, which corresponds to the middle of our spectral interval. The scatter in the spots is considerably larger than at the edge, primarily due to field curvature. The optimal focus position is one where the rings are in focus at the edge of the field (where better imaging is required), and thus the rings nearer the center are out of focus. The RMS scatter in these points is $48 \mu\text{m}$. This should be compared with the $141 \mu\text{m}$ width of an 8 km s^{-1} spectral interval at this radius on the chip.

The large splotch in the lower part of Figure 3.14 shows the spot diagram resulting from on-axis rays. The scatter here is even larger and again is primarily due to field curvature. The RMS scatter of these points is $127 \mu\text{m}$, but the imaging requirements in the center of the ring pattern are very relaxed; the central 8 km s^{-1} spectral element is a disk with a radius of 1 mm.

3.3.2 Sky Imaging Performance

By inserting three lenses into the beam between the etalons and the filter wheel, the WHAM optical system is capable of imaging the sky onto the CCD chip. Fig-

ure 3.15 shows a contour plot of an image of the star γ Gem. The coordinates of the axes are pixel units and each pixel corresponds to $\simeq 0.8'$ on the sky. The bold contour has a level of 1225 ADUs which is half the peak value. Many of the other contours are labeled in ADUs. The star image is quite elongated. The angular size of the portion of the star image within the 500 ADU contour (within which about half of the signal from the star is contained) is $\simeq 5' \times 1.5'$. We believe this elongation is due to a slight warp in one of the siderostat mirrors because the change in the orientation of the image's elongation rotates with the siderostat, but we have not looked carefully into this issue yet due to the low priority of improving the sky imaging performance. Spherical aberration in the large, plano-convex siderostat lens will ultimately limit imaging performance to about $1'$.

3.4 Linearity

A special study was done to test the linearity of the WHAM CCD camera system response at low light levels. The WHAM diffuse light source box that I designed and built in the physics student shop was placed just above the etalons. The light box contained the $H\alpha$ - $D\alpha$ hollow cathode lamp, which is a special lamp containing both hydrogen and deuterium atoms. A stack of neutral density filters were placed in front of the hollow cathode lamp in order to reduce the light level to approximately 1 Rayleigh, a typical high-latitude diffuse galactic background $H\alpha$ intensity. Images were taken at a range of exposure times. For the two lowest exposure times (7.5s and 15s), I averaged multiple exposures (9 and 3, respectively) to keep the signal to noise ratio commensurate with the longer exposure time images.

The resulting spectra are shown in Figure 3.16. The peak on the left, at $\simeq -50$

km s^{-1} , in each spectra is the $\text{D}\alpha$ line (i.e. the Balmer- α line of deuterium), and the peak on the right, at $\simeq +30 \text{ km s}^{-1}$ is the $\text{H}\alpha$ line. The 120s exposure data points are plotted as circles and the remaining spectra are plotted as lines. Each spectra was fitted with two Gaussians and a parabolic background.

Figure 3.17 shows the peak height of the $\text{D}\alpha$ line versus the exposure time. The peak heights were derived from the fitted areas and widths of the Gaussian, and are expressed in units of electrons per 1×1 pixel on the CCD chip. A 1 R source (typical of the intensity of $\text{H}\alpha$ from the WIM at high galactic latitudes) will produce $\simeq 1.6$ electrons (0.75 signal electrons plus the sky background) in pixels at the peak of the line (ring) in a 30 second exposure, which is our intended standard $\text{H}\alpha$ survey exposure time. A typical intensity from the diffuse ionized gas in the galactic plane is about 5 R, corresponding to $\simeq 5$ electrons per 1×1 pixel in a 30s exposure. Figure 3.17 shows that the the CCD response is linear (to within the errors of the study) over this range of charge levels.

The fainter $\text{H}\alpha$ line allows us to test for linearity at even lower light levels. Figure 3.18 shows the peak height of the $\text{H}\alpha$ line versus exposure time. This graph shows that a linear relationship exists for charge levels down to approximately $0.2 \text{ e}^- / 1 \times 1$ pixel, well below the levels that will be typical in WHAM survey images.

3.5 Noise

The primary reason that the WHAM spectrometer is an improvement over the single channel PMT based spectrometer is that it is more efficient. This efficiency translates into the ability to measure a spectrum of a given signal to noise ratio (S/N) in a much shorter time. In this section I discuss the sources of noise in the WHAM

spectrometer and predict S/N ratios for a range of conditions.

First, I derive a general formula for the expected S/N ratio as a function of the exposure time, t . I calculate the S/N obtained for a single 4×4 pixel, since this is the most straight forward number to calculate and all other results can be simply calculated from this. The two types of noise to consider in the WHAM spectrometer are Poisson noise on the number of electrons accumulated in a pixel and the read noise associated with determining that number. The shot noise term depends on the number of accumulated electrons, which depends on the exposure time. In addition to the signal (emission line) generated electrons, there are background generated electrons which are a combination of dark current and electrons generated from sky continuum photons. The read noise does not depend on the exposure time and is a simple fixed number for a given binning factor and gain. The different sources of noise are independent and thus add in quadrature. Therefore, the signal to noise ratio in time t can be calculated from

$$\frac{S}{N} = \frac{\dot{S}t}{\sqrt{(\dot{S} + \dot{K} + \dot{D})t + R^2}} \quad (3.1)$$

where $\dot{S}$ is the rate for generation of signal electrons, t is the time, $\dot{K}$ is the rate for sky photon generated electrons, $\dot{D}$ is the dark current, and R is the read noise. The rates are all in units of electrons per second per 4×4 pixel.

By plugging in the numbers appropriate for our instrument, the expected S/N can be calculated for various cases of interest. For a 1 R source, $\dot{S} = 0.40 \text{ e}^-/\text{s}$ at line center for a line of width 27 km s^{-1} (NAN was used for this calibration). In Madison, the sky rate $\dot{K} \simeq 0.95 \text{ e}^-/\text{s}$ and at Kitt Peak we expect the sky rate to be about half this value. The read noise for 4×4 binning and operating at a gain of $4 \times$

is 4.0 e^- per pixel. Figure 3.19 shows the S/N per data point as a function of time for a 1 R line at line center (assuming a line width of 27 km s^{-1}). This is a typical intensity for $\text{H}\alpha$ from the diffuse ionized hydrogen at high Galactic latitudes. The solid line corresponds to the Madison sky brightness and the dashed curve is based on the estimated sky brightness at Kitt Peak ($0.48 \text{ e}^-/\text{s}$ per pixel). Note that there are 85 pixels in each annulus from which a data point is calculated, and thus the results from Equation 3.1 need to be multiplied by $\sqrt{85}$ to calculate the S/N per data point. At small exposure times, the S/N increases linearly with time as can be seen from Equation 3.1 by neglecting the time dependent term in the denominator. At large times the read noise becomes negligible and the curve approaches the familiar shot noise $\sqrt{t}$ dependence. It must be emphasized that this curve is the S/N per 2 km s^{-1} data point, and thus the S/N per 8 km s^{-1} will be twice these values. The WHAM survey observations will use 30 second exposure times, leading to a signal to noise ratio of ≥ 30 per 8 km s^{-1} spectral element for typical high latitude WIM components.

Figure 3.20 shows the S/N per data point expected for a 0.1 R line. Again the solid curve is appropriate for the Madison sky brightness, and the dotted curve is for Kitt Peak. Notice that when attempting to go very deep by using long exposures, the night sky brightness becomes very important. Also note that in a 15 minute exposure, WHAM can make a statistically solid detection ($S/N \geq 20$ per 8 km s^{-1} spectral element) of a 0.1 R line. In fact, for this very faint work, low level night sky features confuse matters, and the statistical S/N is not necessarily the primary concern. Figure 3.21 shows the S/N that is expected for very long exposures of a 0.01 R line. If the statistical noise were the primary source of errors, one could expect good measurements of emissions at this very low level in several hours. In reality, measurements of

emission at this level will require a good knowledge of the sky background spectrum, which in many cases varies with time. Therefore, such spectra will usually be limited by baseline uncertainties instead of statistical uncertainties. Examples of emission lines with expected intensities at the 0.01 R level include diagnostic lines such as [O I] $\lambda 6300$ and He I $\lambda 5876$ at high galactic latitude, and the $H\alpha$ surface brightness of H I clouds illuminated by the quasar background.

The following is a sample S/N calculation aimed at comparing the achieved S/N to the above calculations. Figure 3.22 shows the $H\alpha$ spectrum of a direction towards the high velocity cloud (HVC) Complex M (see Chapter 4). This spectrum is the average of five separate 15 minute exposures taken “on” the HVC direction minus 5 paired “off” directions taken in nearby directions. The measured intensity of the line is 0.077 R and the line width is $\simeq 30 \text{ km s}^{-1}$ (FWHM). The read noise is negligible compared to the Poisson noise on the sky counts in this case, and therefore the expected S/N can be calculated by plugging the total exposure time (9000 seconds) into Equation 3.1 and dividing the result by two since the signal is only present during half the exposure time. The result is $S/N = 13.6$ per data point. The σ per data point calculated from the scatter in the residuals shown in the lower part of Figure 3.22 is 4.6 ADUs and the peak height of the signal is approximately 45 ADUs. Thus the measured S/N is 9.8. This is slightly lower than the predicted S/N , and the difference appears to be due to structure in the baseline that was not allowed for in the one parameter baseline fit. In an attempt to measure the true scatter in the residuals, I fit a line to the residual data points with $v_{lsr} > -115 \text{ km s}^{-1}$ (the right half of the range shown in Figure 3.22) to account for the upward slope seen in the data and measured the scatter about this line to be $\sigma = 3.6$. This noise level would imply a $S/N = 12.5$,

quite close to the predicted S/N .

3.6 WHAM Efficiency Compared to Single Channel (PMT) Mode

The previous Wisconsin large-aperture Fabry-Perot spectrometer, used in many ground breaking studies of the WIM, used a single-channel photo-multiplier tube (PMT) for its detector. By utilizing a modern CCD detector, the WHAM spectrometer can achieve large gains in efficiency over the PMT based detector (See Section 2.3.4).

We did a special test to compare directly the efficiency of the WHAM spectrometer to the older, PMT based spectrometer. The $H\alpha$ - $D\alpha$ hollow-cathode lamp was placed into the WHAM diffuse light source box, and the intensity of the $H\alpha$ line was decreased to about 1 R by placing neutral density filters in front of the lamp. We calibrated the intensity of the box using both a “standard” $H\alpha$ source used for aeronomy measurements, and using the old spectrometer (the sensitivity of which is known from nebular calibrations). Before disassembling the old spectrometer, this light source was placed above the etalons and a 1090 second scan was taken that covered an approximately 200 km s^{-1} radial velocity interval around the lines. The resulting spectrum, shown in Figure 3.23, contains the Balmer-alpha line of hydrogen ($H\alpha$) on the right, and the Balmer-alpha line of deuterium ($D\alpha$) on the left. The integration time per data point is 10 seconds.

Figure 3.24 shows the spectrum of the same $\simeq 1 \text{ R}$ diffuse light source measured in a 15 second exposure with the WHAM spectrometer. The signal to noise for the $H\alpha$ line of this spectrum is $\simeq 17$, where the signal to noise of the $H\alpha$ line calculated

for the spectrum measured with the old spectrometer in 1090 seconds is $\simeq 13$. The gain in efficiency is thus better than a factor of 70! The gain over the PMT based spectrometer will be even larger for longer exposures. Because it uses a CCD detector, the WHAM spectrometer noise contains a read noise term which is not present in the old spectrometer. The read noise term does not depend on the exposure time, and hence as the exposure time is increased, the read noise becomes less significant relative to the other, time dependent sources of noise. Since 15 seconds is the shortest exposure time we will use with the WHAM spectrometer, the above factor of 70 is a lower limit to the gains achieved with the WHAM spectrometer. Gains are more than 100 fold for exposures that are long enough to render the read noise term unimportant. This factor results from the approximately $4\times$ gain in quantum efficiency combined with the gain that results from the fact that the WHAM spectrometer samples the 25 spectral elements simultaneously, while the PMT based spectrometer was only sensitive to one spectral element at a time. Of course, it must be remembered that at the higher signal to noise levels achieved with WHAM, baseline uncertainty associated with the the night sky background is the ultimate limitation.

It is possible to compare the photon detection efficiency of the PMT based system to the WHAM system using these spectra. The WHAM spectrum is plotted with the number of ADUs per pixel, averaged over the ring, on the ordinate. Each data point in the plotted spectrum corresponds to a 2 km s^{-1} interval, which derives from an annular ring on the CCD chip containing 85 pixels. Hence, the number of ADUs measured at the peak of the $\text{H}\alpha$ line is $7 \times 85 = 595$, which corresponds to 340 electrons, or 23 electrons per second of exposure time. For the PMT scan, the $\text{H}\alpha$ line has 108 electrons at its peak, or about 11 per second of exposure time. In this

case, however, the aperture passes a full 8 km s^{-1} spectral interval on the detector. Therefore, to make a fair comparison with the 2 km s^{-1} interval used for each data point in the WHAM spectrometer, we must divide by $\simeq 3$ (allowing for the fall off from the peak in the line). This leads to 3.7 electrons per 2 km s^{-1} interval implying that the WHAM spectrometer is detecting approximately 6 times more photons per wavelength interval per time. The increase due to the relative quantum efficiencies of the two detectors is about a factor of 4, and the additional factor of 1.5 in the gain is probably due to the fact that nearly all of the optical elements of the WHAM spectrometer have multilayer anti-reflection coatings with a reflectivity of $\lesssim 0.25\%$ per air-glass surface, while the old spectrometer used single layer MgF coatings (1.5% per surface). Another contributing factor is a vignetting loss associated with passing the light onto the photocathode of the PMT in the old system.

3.7 Intensity Calibration

The intensity calibration for the WHAM spectrometer is determined using astronomical sources of known intensity. The current intensity calibration is based on the work of Scherb (1981) who measured the absolute intensity of NGC 7000 (also called the North American Nebula, or NAN). Scherb's calibration of NAN used Fabry-Perot measurements and relied on the planetary nebula NGC 7662, and the stars α Boo and α Tau for reference objects. The result is that the $\text{H}\alpha$ surface brightness for NAN within a $49'$ beam (centered at R.A.(1950) = $20^{\text{h}} 56^{\text{m}} 17^{\text{s}}$, dec(1950) = $44^\circ 24' 3''$) is 850 ± 50 Rayleighs. This forms the primary calibration source, and secondary sources are calibrated from this for use during the times of year when NAN is not available for observation. During every observing night spectra are taken of NAN or one of

the secondary calibration sources to allow correction for night to night variations in atmospheric transmission. The intensities of all other observations for the night are then derived by comparison to the calibration source spectrum, including a zenith angle correction.

Using Scherb's calibration introduces some systematic errors into our intensity calibration. One problem is that the above $H\alpha$ intensity calibration applies to a $49'$ beam, while the WHAM beam is 1° in diameter. Since the intensity of the nebula falls off with distance from the point of maximum intensity, the average over our larger beam will be slightly smaller than for the beam used in the calibration. A related problem is that both the spectrometer used for the calibration and the WHAM spectrometer are not uniformly sensitive within their beams. Hence, what is really needed is the average brightness of the calibration nebula within the spectrometer beam, weighted by the beam sensitivity function. We will use a calibrated, higher spatial resolution $H\alpha$ image of NAN to calculate the average surface brightness of NAN within our beam, weighted by the beam sensitivity function.

In the near future we will carry out a study to improve the accuracy of the WHAM intensity calibration. One important step in the process is to measure the beam sensitivity function. A bright star provides a steady point source of light. By taking spectra with the star placed at a grid of locations within and near the beam, we can map out the spectrometer's sensitivity as a function of location with respect to the beam center. We are able to use the WHAM mapping software, developed by Matt Haffner, to place the star at a grid of locations within the beam by using a grid spacing of $\simeq 0.1^\circ$, instead of the usual 1° . We carried out a first attempt at this measurement using α Boo as the calibration star on a night with non-optimal

sky conditions and the resulting spectra are shown in Figure 3.25. A grid of spectra are shown and the location of a spectra on the figure corresponds to the offset of the look direction from one where the star is centered in the beam. A circular pattern is apparent in the figure, showing which spectra have the star within the beam. Each spectrum within the beam contains the stellar continuum near $H\alpha$. On the right side of each of these spectra, the star's $H\alpha$ Fraunhofer absorption trough is apparent with the geocoronal $H\alpha$ emission line just to the blue of the minimum. Near the center of the beam one can see several spectra that were affected by a cloud drifting by. The geocoronal emission arises from a diffuse source and is seen to be less attenuated than the light from the star in the cloud contaminated spectra. Another problem with this preliminary attempt at the beam sensitivity map is that the grid of directions covers too small a region of the sky and therefore the beam was not fully covered, as can be seen on the right and at the bottom of the figure. Also, higher angular resolution sampling is needed to better define the response pattern near the edge of the beam. By measuring the average continuum level for each of the spectra, we can map out the beam sensitivity function.

Another important property of the WHAM system that we will soon investigate is the sensitivity as a function of wavelength. We intend to determine this by measuring the spectrum of a well calibrated star at a range of wavelengths. Using the spectrophotometric stellar spectra collected by Breger (1976) along with the high resolution spectral atlases of Griffin (e.g., Griffin 1968), which exist for α Boo and α CMi, we can determine the absolute brightness of the stellar emission that falls within our bandpass at a series of wavelengths. Combining these with our measured spectra will provide a measurement of the WHAM spectrometer's relative sensitivity

to different wavelengths.

3.8 CCD Characteristics

In general terms, the CCD camera has three functions. Consider light incident on a single pixel of the chip. The incoming photons must be transformed into electrons that are stored at the pixel location. When the exposure is completed, the charges at this pixel must be transferred pixel by pixel to the on-chip amplifier. Finally, the amplifier must ‘read’ the amount of charge in the charge packet. This last step requires the charge packet to be transformed into a voltage, amplified, and then digitized for storage. In this section I will summarize the properties of the WHAM CCD chip, as well as its measured performance characteristics.

3.8.1 The CCD Camera System

The WHAM spectrometer uses a state of the art CCD camera purchased from Photometrics Ltd for its imaging detector. Figure 3.26 shows a schematic overview of our camera system. The CCD chip itself was manufactured by Tektronix, and its properties and performance are summarized in the next section. The chip must be kept cold in order to minimize thermally generated dark current. For this reason, the chip is housed in a cryogenic dewar containing liquid nitrogen. The camera is run by a camera controller, also provided by Photometrics, which contains electronics for driving the chip, along with devices that regulate the temperature of the chip to within $\simeq 0.1^\circ\text{C}$. The controller also initiates opening and shutting the shutter, which is placed just upstream from the camera lens (see Figure 3.26).

The WHAM workstation is command central for all WHAM spectrometer ac-

tions. It is interfaced through a terminal server to all of the WHAM control subsystems including, for example, the PC, which commands the CCD camera, and the instrument control computer chip, which controls the auto-fill valve. The workstation sends camera commands to the PC over a serial line and receives data back from the PC via an ethernet connection. The IBM compatible PC is interfaced to the Photometrics camera controller. Photometrics provides a library of C functions that run on the PC to allow programmable control of the camera. John Lehmke of the University of Wisconsin Space Astronomy Lab (SAL) wrote a program for the PC that accepts high-level commands over a serial line from the workstation, and then sets up and calls the appropriate C control functions to carry out the desired actions. This program also retrieves and stores the images, and then transfers them via ethernet to the workstation disk. From there, software running on the workstation processes the images into spectra and displays both the images and the spectra on the workstation screen.

Kurt Jaehnig of SAL has developed a liquid nitrogen auto-fill system to allow remote filling of the camera dewar. A large (160 liter) dewar sits outside the WHAM trailer and has a vacuum insulated flexible line leading to the CCD through an auto-fill valve. This valve is controlled by the WHAM instrument control computer which is in turn commandable from the workstation.

3.8.2 General Properties of the CCD Chip

The CCD chip chosen for the WHAM camera system is the Tektronix 1024AB2, which contains 1024×1024 pixels $24 \mu\text{m}$ on a side. The chip is thinned and back-illuminated and utilizes multi-pinned phase (MPP) mode to minimize dark current (see

below). This chip was chosen based on its low read noise, high quantum efficiency, and low dark rate. A further consideration was the chip's ability to maintain a high charge transfer efficiency at very low charge levels. The chip is read out at 16 bit resolution with a 40 kHz clock rate. It is possible to operate at different gain factors which determine the number of arbitrary data units (ADUs) assigned to each electron in the digitization process. For higher gain factors, the system achieves lower read noise at the cost of longer readout time. For gain = 1, there are 2.18 electrons per ADU, the read noise is 4.62 e⁻ RMS, and the usable full well capacity (limited not by the actual depth of the well but instead by the analog to digital converter) is 142 ke⁻. At gain = 4, there are 0.57 electrons per ADU, the read noise is 2.83 e⁻, and the usable well capacity is $\simeq 37$ ke⁻. Because of the lower read noise, and because a signal of 37 ke⁻ is never reached in normal operation, we use gain = 4.

3.8.3 Quantum Efficiency

The efficiency with which the incoming photons are transformed into electrons is called the quantum efficiency (QE). In a normal front side illuminated CCD chip, the quantum efficiency is hindered by the overlying gate structure that is necessary to hold charge at a pixel and to transfer charge. To achieve the highest possible quantum efficiency, it is necessary to illuminate the chip from the back. This further requires the chip to be thinned, which is accomplished by chemically etching material off the back of the chip. The Tektronix chip employed by WHAM is of the back-illuminated, thinned variety with an antireflection coating to achieve the maximum possible QE. The QE depends on the wavelength, and a graph of QE versus wavelength for the WHAM chip, which was measured by Photometrics before shipment, is shown in

Figure 3.27. Near $H\alpha$ the QE is 78%, about 4 times the QE for the PMT used in the previous spectrometer.

3.8.4 Charge Transfer Efficiency

The charge transfer efficiency (CTE) is defined as the average fraction of the charge in a pixel that is successfully moved to the next pixel in a charge transfer operation. The electrons that are not successfully transferred to the adjacent pixel most often are not lost but remain in the source pixel and are one step behind the rest of the packet in subsequent transfers. This is called deferred charge. There are also charges that are completely lost, but this occurs at a lower rate. Consider the affects of an imperfect CTE on a WHAM spectrum. In the case where the spectrum contains a single emission line, there will be a circular ring of pixels containing enhanced charge levels. In the process of transferring the charge through the parallel registers, the ring will be slightly broadened vertically due to the charge that lags behind. This broadening will be most pronounced at the top of the ring since that is where the most transfers are necessary and also the charge trails perpendicular to the ring tangent at the top. For the serial transfers there is a horizontal smearing of the ring pattern and the right part of the ring is most affected. If all the unsuccessful transfers lead to deferred charge (no charge is lost), then the intensity of the line (derived from the area of the spectral feature) will be accurately measured. The measured emission line will be somewhat broadened, however, due the the charge smearing discussed above. If the CTE were as low as 0.999, the deferred charge would lead to a smearing of about one unbinned pixel, which corresponds to approximately 2 km s^{-1} at the extreme edge of the spectrum. However, the distortion in the actual line profile in the

resulting spectrum would be at a much lower level because this amount of smearing would be limited to only a small portion of the ring. It is much more problematic if the charges that are unsuccessfully transferred are lost. In this case, since the CTE decreases with decreasing charge levels, there will be a nonlinearity in the system. Unfortunately, it is very difficult to find information in the literature on the fraction of the unsuccessful transfers that arise from lost charges for low charge levels. For this reason, we performed a linearity test to place a limit on this effect (see below). A study we may carry out in the future is to measure the line profile of a faint thorium line versus the integration time to look for the charge smearing effects described above.

The TK1024 chip is capable of a CTE of 0.99999 at high charge levels, but the CTE generally decreases with decreasing charge level. This is of special concern for the WHAM spectrometer since we are interested in studying extremely faint sources. For a source with an intensity of 1 R (typical of the diffuse Galactic H α intensity at high galactic latitudes), a 30 second exposure with the WHAM spectrometer will lead to the generation of $\simeq 1.64$ electrons of charge per 1×1 pixel (counting the expected sky background rate at the Kitt Peak site) at line center. Figure 3.28 shows measurements of the charge transfer efficiency of a TK1024 CCD chip that was under development for use in the STIS and SOHO instruments (Orbock et al 1990). The triangles in the figure show that $(1 - \text{CTE})$ increases as the charge level per pixel is decreased. The lowest measured charge level is 10 electrons per pixel, where the CTE is measured to be 0.999245, significantly higher than the charge levels we are interested in for the WHAM spectrometer; however, the trend in the curve suggests that the CTE may level off around 0.999 for very low charge levels. We had Photometrics do a special test of the charge transfer efficiency at low charge levels before they shipped the camera.

The results of their study suggest that the CTE is ≥ 0.999 at a charge level of 0.775 electrons per 1×1 pixel, but the study was not capable of distinguishing lost charge from deferred charge.

One of the main motivations for the linearity study presented in Section 3.4 was to verify that the lost charge component of the unsuccessful transfers for our CCD chip is sufficiently small. This section shows that there are no significant nonlinearities in our system down to very low charge levels. Using, for example, the fact that the 30 second $H\alpha$ emission line in this study shows no deviation from linearity and is measured with $\simeq 10\%$ uncertainty, we can derive that the lost charge component of the transfer inefficiency is less than 1 part in 10^4 . The charge level for this measurement is $\simeq 0.8$ electrons per 1×1 pixel; the limit on the lost charge component of the charge transfer inefficiency is indicated as a downward arrow in Figure 3.28.

3.8.5 The Stability of the Bias Level

A bias frame is a dark exposure with zero exposure time. The pixels will not contain zeros in this case, but some offset that is referred to as the bias level. A bias frame is read out in the same manner as any other exposure and provides an opportunity to measure the read noise alone, without the added shot noise of signal counts.

Figure 3.29 shows the bias levels as a function of time using bias frames from the dark current versus temperature study presented below. While the temperature was held constant the bias level remained quite stable. The portions of the curve that drift slowly downward show the bias drift. The temperature of the CCD chip was adjusted several times during this study, and there is a slight dependence of the bias

level on temperature which can be seen in the figure as the steeper steady shifts of the bias level. In several other studies done at constant temperature, I have found that the bias level drifts at ~ 1 electron per hour, but the shape of the bias image remains the same. Since the shape stays the same, the bias level drift only causes a dc offset in our spectra, which becomes a harmless part of the background.

3.8.6 The Read Noise and Spurious Charge Noise

The error in determining the number of electrons in each pixel of the CCD chip is called the read noise. In practice, the read noise has a fixed size, independent of the number of electrons in the pixel. The read noise is an important source of noise in WHAM spectra, especially for short exposures (see Section 3.5). The read noise of the TK1024 chip was measured by Photometrics to be 2.83 electrons RMS when the camera was operated at gain of $4\times$. While in the process of verifying this number, I found that I was able reproduce this number for images using 1×1 binning, but also found that there was an extra source of noise in images with larger binning factors. This unexpected effect is shown in Figure 3.30, where I have plotted the measured value of σ in bias frames taken with various binning factors ranging from 1×1 binning to 11×11 binning. This behavior is unfortunate because we plan to use 4×4 binning for the WHAM ring images, and the noise level is 3.87 electrons per pixel at this binning factor, significantly higher than the 2.73 electrons per pixel measured for the 1×1 (i.e., unbinned) image. We have tentatively identified spurious charge generation as the source of this extra noise. Spurious charge generation is a well known problem associated with CCD chips operating in MPP mode and results from a mechanism known as impact ionization which occurs during charge transfers. I will next present

the results of several studies that I have done to characterize the properties of this unexpected extra noise source.

Histograms of the pixel values in the bias frames at various binning factors show that the extra noise is approximately Gaussian, indicating that the extra noise can be characterized as shot noise, although by the time one gets to 11×11 binning the distribution of values looks somewhat triangular. I have attempted to model this noise by assuming that the read noise is fixed and equal to 2.73 and that the measured noise is the sum of the read noise and the spurious charge noise. In Figure 3.31, I have plotted the $(\text{total noise})^2 - (\text{read noise})^2$ versus the square of the binning factor. The ordinate of this plot is thus the $(\text{extra noise})^2$ and interestingly, it is linear with $(\text{binning factor})^2$; that is, it is linear with the number of 1×1 pixels in a binned pixel. The slope of the fitted line shown in the figure is 0.453 electrons/ $(1 \times 1 \text{ pixel})$. Assuming the extra noise is Poisson noise on spuriously generated electrons, this indicates an extra 0.5 electrons in each 1×1 pixel (the variance is equal to the mean in a Poisson distribution).

In order to determine whether the extra noise is generated preferentially through the serial binning process or the parallel binning process, I took bias frames with rectangular bins, which involved either only serial binning (i.e., $1 \times b$ bins, where b is the binning factor) or only parallel binning (i.e., $b \times 1$ bins). Figure 3.32 and Figure 3.33 show the results for the serial binning and parallel binning cases, respectively. The results are plotted in a similar format to Figure 3.31 with the read noise subtracted from the measured noise on the ordinate. Again, the extra noise is linear with the number of 1×1 pixels in a binned pixel in both the serial binning and parallel binning cases. The slope of the line for the serial binning case is 0.80, and in the parallel

binning case it is 0.41, indicating that the serial transfers have significantly greater spurious charge generation than parallel transfers.

If the extra noise is the result of charge transfers, one might expect the region of the chip requiring more transfers on the way to the amplifier to have more noise. However, Figure 3.34 shows that this is not the case. The figure shows the increase in the measured σ versus binning factor (square bins) for the four quadrants of full frame bias images. The triangles correspond to noise measurements in the lower left quadrant, the squares to the lower right quadrant, the X symbols to the upper right quadrant, and the circles to the upper left quadrant. The amplifier is in the lower left, parallel transfers are top to bottom, and serial transfers are right to left. The pixels in the upper right quadrant of the chip have undergone on average 3 times more charge transfers than the pixels in the lower left quadrant on their way to the amplifier, and thus one might naively expect them to have a much higher spurious charge noise. In fact, if one considers the clear operation that occurred just before the bias exposure was taken, then the number of transfers required is the same for every pixel on the chip. Each pixel is transferred a total of $\simeq 1000$ times and hence the spurious charge generation mechanism must be generating on average an extra electron for every 2000 transfers.

3.8.7 The Dark Current

Charge accumulates in the pixels of the CCD even when the chip is not exposed to light. This “dark current” is the result of electrons that are thermally excited across the silicon band gap and thus depends strongly on the temperature. The main concern with the dark current is the additional noise that results from Poisson statistics on the

dark counts. The simple solution is to cool the CCD chip to a low enough temperature to make the dark current negligible in relation to the other sources of noise. This is why the CCD camera requires a liquid nitrogen cooling system.

At what chip temperature should we operate? The minimization of noise on dark counts would argue for operating at the lowest temperature attainable. However, there are two reasons for wanting a higher temperature: 1) The CTE increases with temperature, and 2) the spurious charge generation is expected to become less severe at higher temperatures. At very low temperatures the dark current falls to completely negligible levels. The recommended strategy is to increase the temperature to the maximum value allowed by our tolerance for noise on the dark counts. For WHAM, this requires keeping the dark noise down to a small fraction of the noise on the sky counts. In Madison, the sky rate is $\simeq 1500$ ADUs in a 4×4 pixel in 15 minutes. At Kitt Peak we expect the sky to be about half this bright. The estimated Kitt Peak site sky rate reduces to $\simeq 0.9$ electrons per 30 seconds per 1×1 pixel. If we want the dark noise to be 10% or less of the noise on the sky counts, then we need to keep the dark current under a third of this rate, or 0.3 electrons per 30 seconds per 1×1 pixel.

I collected a set of images aimed at determining the dark current as a function of temperature. With the camera carefully shielded from all light, bias frames and dark images were taken alternately. At each of four temperatures, dark images were obtained at a range of exposure times to insure that the measured signal was linear with time. The solid lines in Figure 3.35 show the measured number of dark counts versus the exposure time for CCD chip temperatures of -80.6°C , -90.7°C , -100.8°C , and -110.8°C (from top to bottom). The dark counts were determined by averaging the number of counts per pixel in the dark image and subtracting the average number

of counts per pixel in the bias image taken just before the dark image. Special care was taken to eliminate the effects of cosmic ray contaminated pixels on the average values determined. The dark rates are determined from the slopes of the solid lines shown in Figure 3.35.

Figure 3.36 shows the dark counts measured for the $T = -100.8^\circ \text{C}$ case. Unlike the data gathered at the other temperatures, the points do not fall on a straight line but rather show a downward curvature. These data were gathered just after the chip was cooled down from room temperature and we suspect that while the temperature gauge read -100.8°C , the chip had not yet reached this temperature and was still in the process of cooling. The dotted line in Figure 3.35 shows a fit to the first five data points taken at $T = -100.8^\circ \text{C}$, and the solid line just below it shows a line drawn from the origin to the $t = 1200\text{s}$ data point for $T = -100.8^\circ \text{C}$. We will adapt the slope of the solid line as the dark rate for this temperature.

Figure 3.37 shows the dependence of the measured dark current versus chip temperature. The dark currents were determined from the slopes of the solid lines shown in Figure 3.35 and are plotted in units of electrons per 30 seconds per 1×1 pixel (30 seconds is the exposure time that will be used for survey observations). Even for the $T = -80.6^\circ \text{C}$, the rate is significantly lower than the $0.3 \text{ e}^- / 30 \text{ seconds per } 1 \times 1$ pixel upper limit calculated above based on considerations of the tolerable dark noise level.

The dark images show that dark current is not produced uniformly over the chip. Figure 3.38 shows the 600 second exposure dark image for the $T = -80.6^\circ \text{C}$ case. Broad, curved bands are clearly visible in this image. Slices of this image show that the peak to trough ratio of these bands is $\simeq 1.3$. Bands are also visible on the T

= -90.7° C and T = -100.8° C dark images, and these bands show the same peak to trough ratio. No bands are apparent on the T = -110.8° C images, but the dark rate is so low at this temperature that the bands, if present, would be swamped by the noise. The shape, size, and location of the bands looks the same on all the images in which they are seen, arguing against electronic pickup of rf noise as the source of the bands. The same band structure was seen in a 600 second dark image that I took over 1.5 years ago, implying this structure is a stable feature in the chip's dark current. These bands will introduce systematic irregularities into the measured spectra, especially near the center of the ring pattern, where the bands are not averaged out over a ring. The peak to trough difference in the 600 second, T = -80.6° C is about 20 ADUs, which in a 900 second exposure would correspond to the signal produced by a $\simeq 0.05$ R source. At -100.8° C the contamination enters at the 0.01 R level for a 900 second exposure, which is about our noise level for these spectra. Hence, we will operate at -100.8° C; I have shown elsewhere that the CTE is sufficiently high and that the spurious charge generation is manageably low at T = -100.8° C, which has been our standard operating temperature.

3.9 Chromatic Aberration: Best Focus Position versus Wavelength

The effective focal length of the ring imaging lenses is shorter for blue light than it is for red light. This is due primarily to the chromatic aberration of the lens just past the etalons (L3 in Chapter 2, Figure 2.13), and means that for blue light the camera lens must be placed slightly closer to the chip than for red light. To accomplish this,

the WHAM camera lens system is mounted on a movable track, and the exact position of the lens with respect to the chip is adjusted by a micrometer attached to the lens mount.

To determine the best focus position for a given wavelength, it is necessary to find a narrow line, place the ring corresponding to that line near the outer edge of the ring image, and then record the image of that ring for a range of focus positions. A plot of the resulting width of the measured spectral line versus focus position reveals the best focus position, that is, the micrometer position for which the line is narrowest. The line must be placed at the outer (blue) edge of the ring image because that is where the rings are physically the narrowest and therefore where it is important to have optimal imaging (see Chapter 2, Section 2.3.5). We used various lines emitted by the thorium atom excited by a hollow-cathode lamp. Hollow-cathode lamps are designed to excite the emitting atoms without raising them to high temperature, and thorium atoms are very heavy and thus have a smaller thermal broadening at a given temperature than lighter atoms. Another advantage of the thorium atoms is that they emit many lines with wavelengths that cover the whole spectral range. This allows us to use our existing filters and find thorium lines that lie within the bandpass of these filters. For this particular study, I used a thorium line near $H\alpha$, a thorium line near $H\beta$, and one near the $\lambda 5755$ line of ionized nitrogen.

Figure 3.39 shows a spectra of the thorium $\lambda 5753$ line that was passed by our N II $\lambda 5755$ filter with the focus set near the optimal value. The rise at the left edge of this spectrum is due to a light leak problem that this filter has near the edge of the aperture (our other filters do not have this problem). The fit was performed using two Gaussians, a broad one and a narrow one. The broad component was fixed in the

fits to spectra at the different focus settings, and Figure 3.40 shows the width of the narrow component versus the focus setting. Figures 3.41 and 3.42 show similar graphs for thorium lines near $H\beta$ ($\lambda 4863$) and $H\alpha$ ($\lambda 6554$), respectively.

Using Code V, the sophisticated ray-tracing that was used to optimize the design for the WHAM optical system, John Harlander of St. Cloud State University modeled the chromatic aberration of the WHAM system. This modeling predicted focus position offsets with respect to $H\alpha$ for a range of wavelengths between $H\beta$ and the [S II] $\lambda 6716$ line. I fitted these points to a cubic polynomial, and the result is plotted as the smooth curve in Figure 3.43. The measured optimal focus positions are also plotted on this graph. The curve and the data points agree to within the experimental error of $\pm 0.03\text{mm}$. Given the good agreement between these two curves, we can now use the model curve to determine the best focus position for an arbitrary wavelength. The uncertainty in the best focus position is sufficiently small so that no significant broadening of the instrumental response profile will result.

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Orbock, J. D., Murata-Seawalt, D., & Delamere, W. A. 1990, SPIE, Vol. 1242, 105

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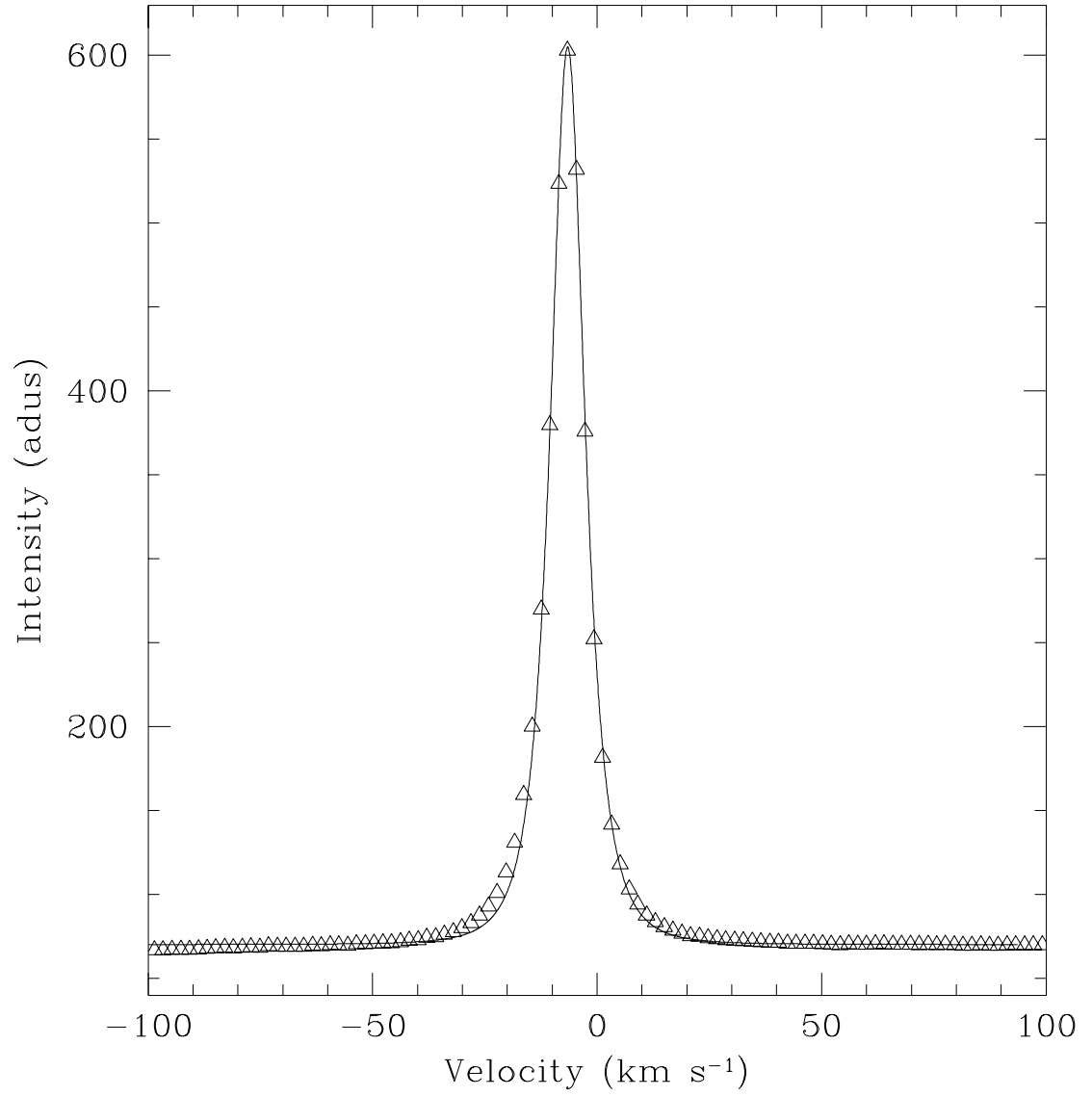


Fig. 3.1.— The instrument profile. The data points show the measured instrument profile and the curve shows a model instrument profile.

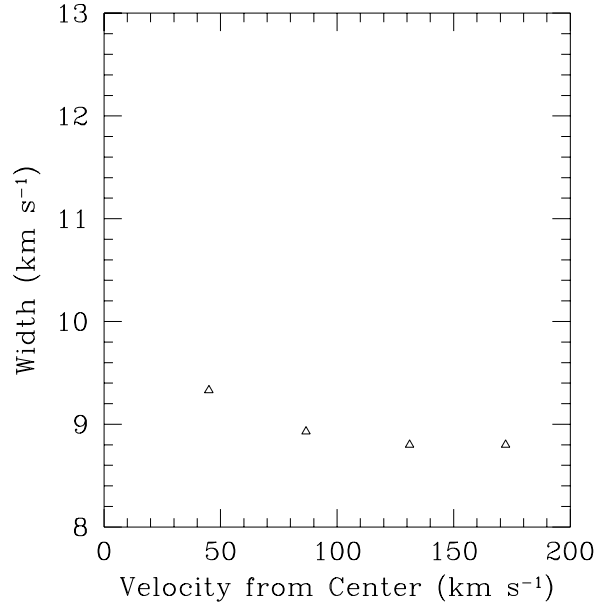


Fig. 3.2.— Instrument profile width versus velocity from the ring pattern center for the case of 1×1 binning with masked etalons.

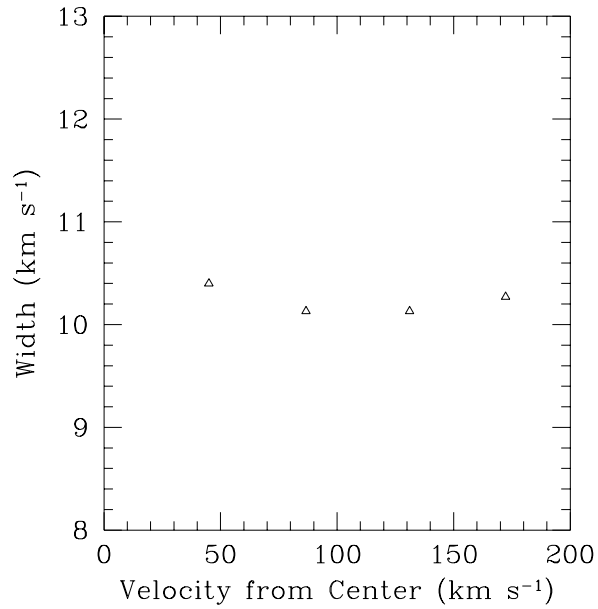


Fig. 3.3.— Instrument profile width versus velocity from the ring pattern center for the case of 1×1 binning with unmasked etalons.

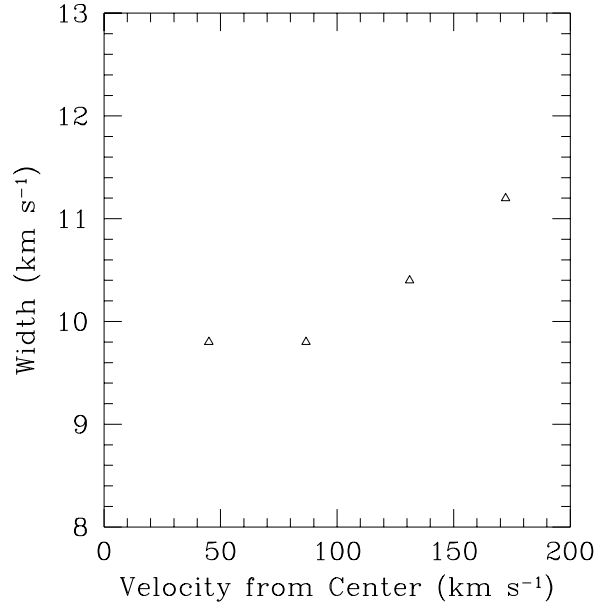


Fig. 3.4.— Instrument profile width versus velocity from the ring pattern center for the case of 4×4 binning with masked etalons.

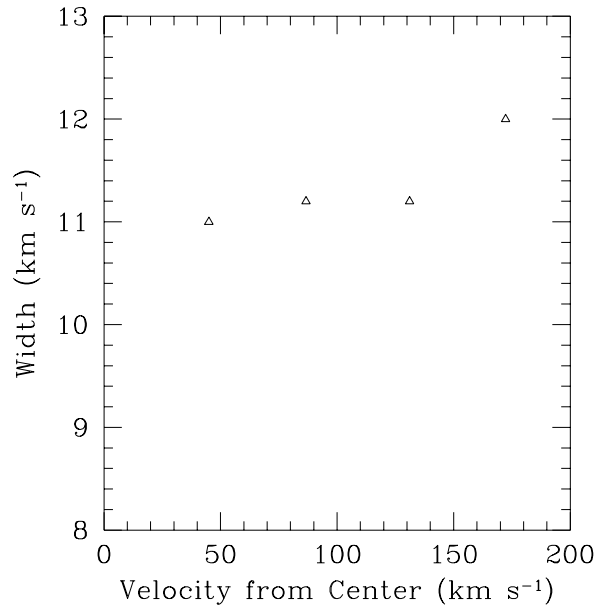


Fig. 3.5.— Instrument profile width versus velocity from the ring pattern center for the case of 4×4 binning with unmasked etalons.

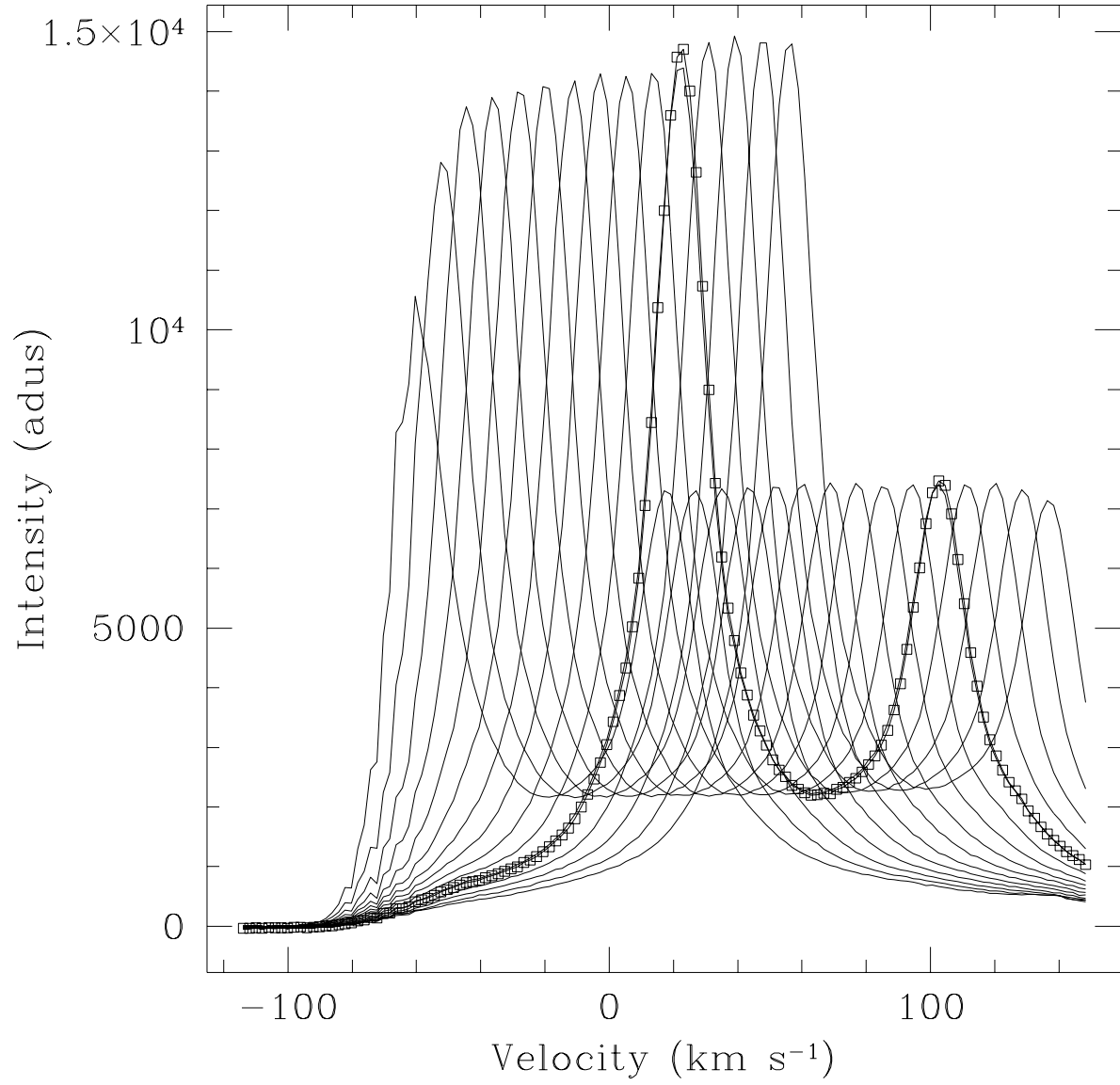


Fig. 3.6.— Spectra of the H α -D α lamp.

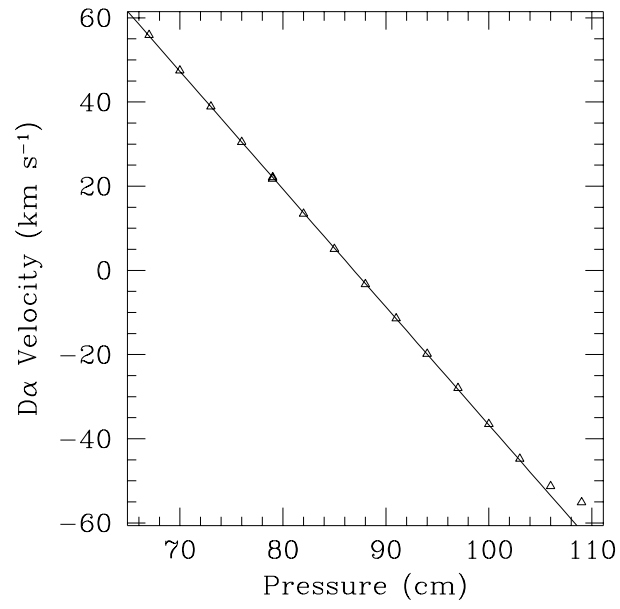


Fig. 3.7.— D α mean velocity vs. pressure.

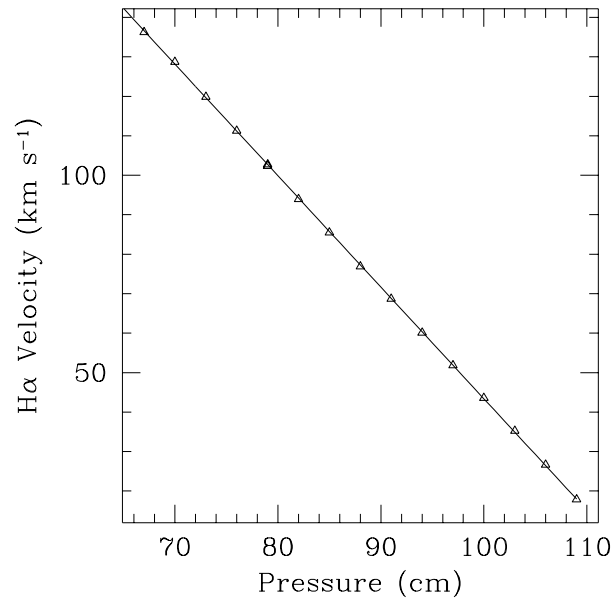


Fig. 3.8.— H α mean velocity vs. pressure.

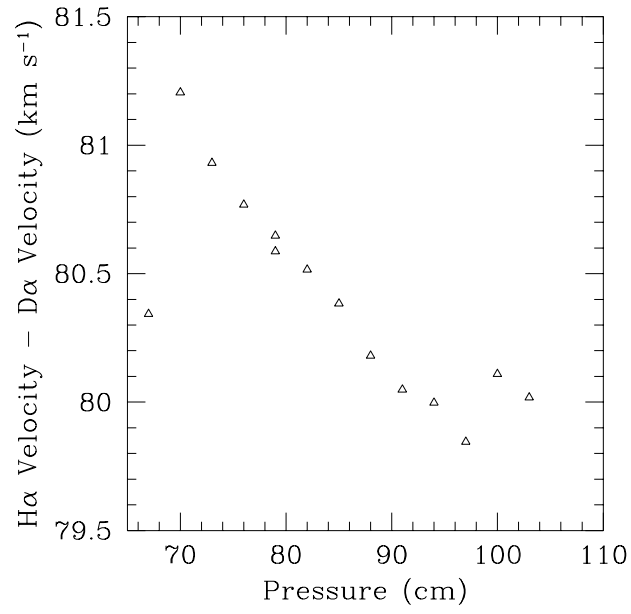


Fig. 3.9.— H α -D α velocity difference vs. pressure.

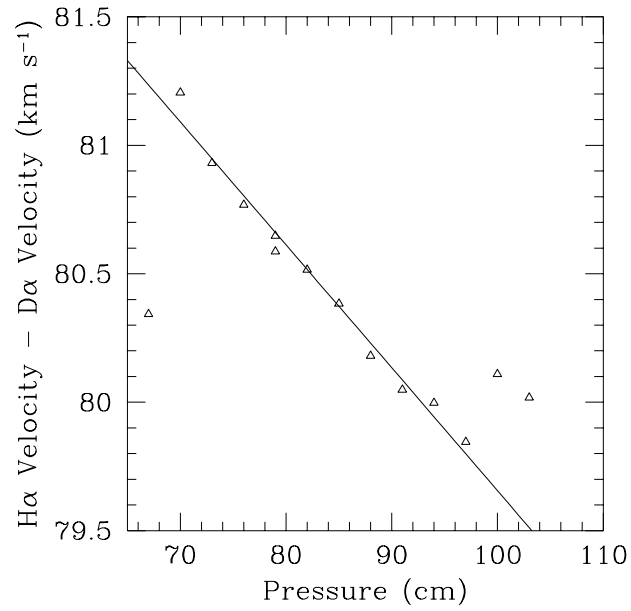


Fig. 3.10.— H α -D α velocity difference vs. pressure, with fit.

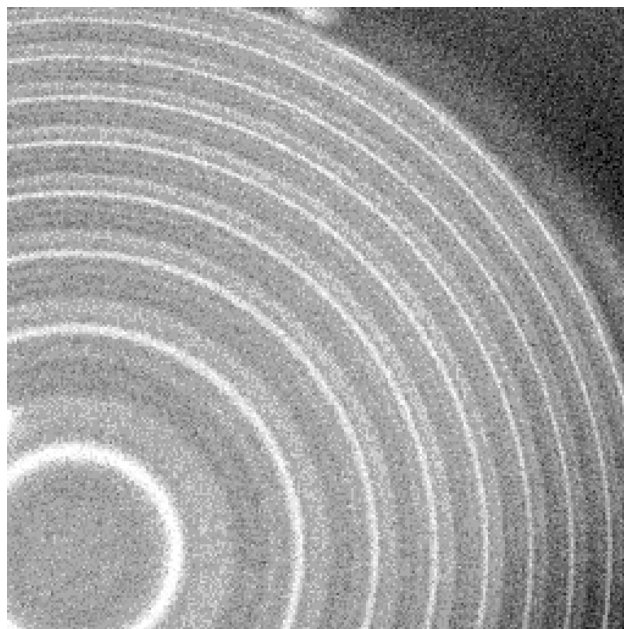


Fig. 3.11.— Thorium rings from very high resolution etalon.

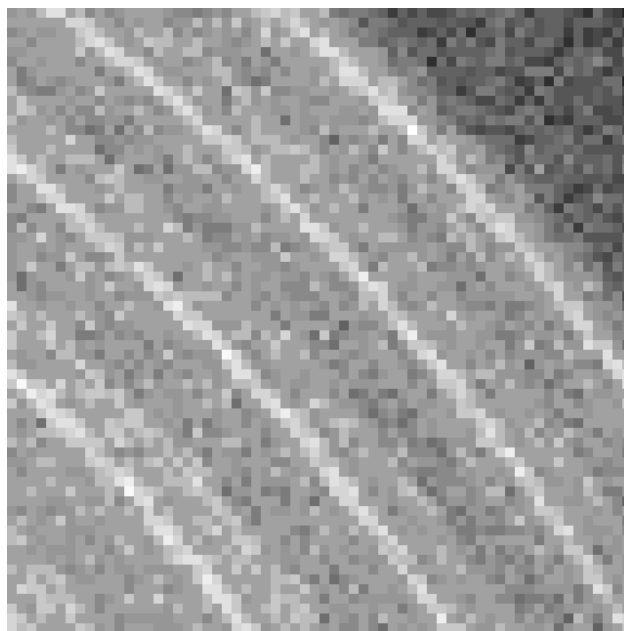


Fig. 3.12.— Close up view of thorium rings from very high resolution etalon.

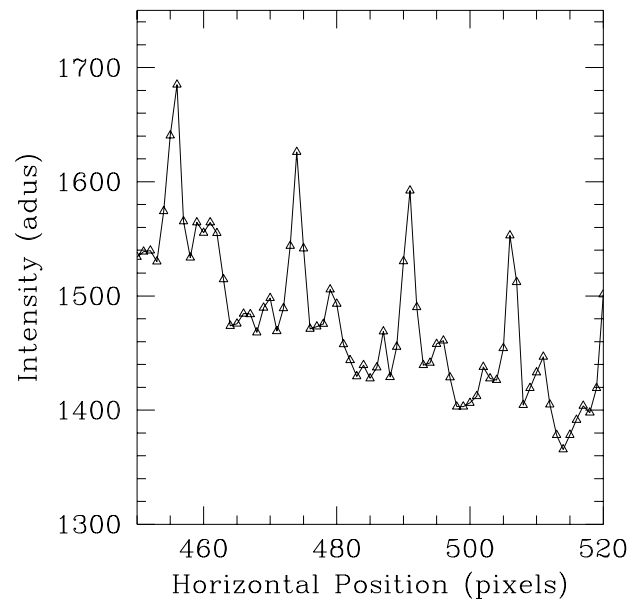


Fig. 3.13.— A horizontal slice from a portion of the very high resolution thorium ring.

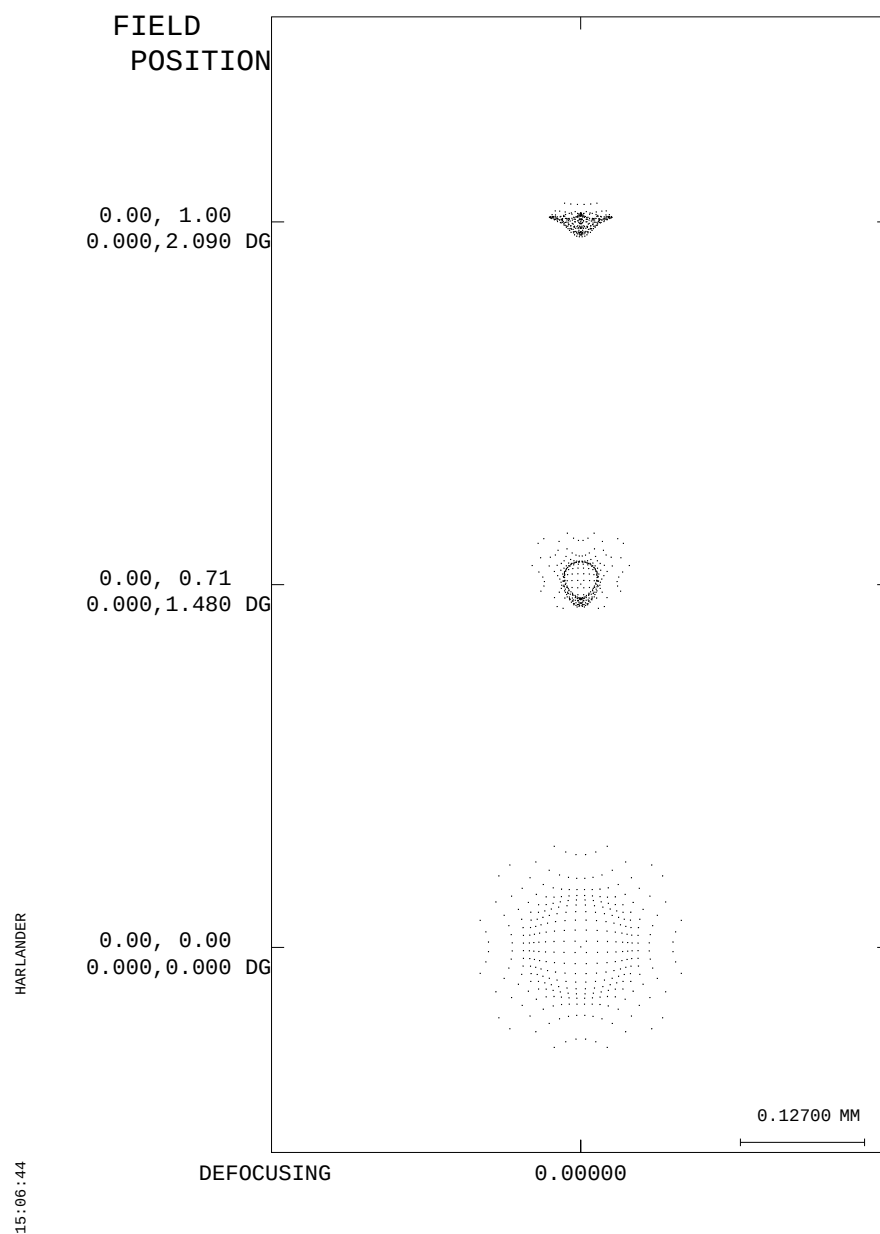


Fig. 3.14.— WHAM ring imaging system spot diagrams for on-axis rays, rays at 1.48° off axis, and rays at 2.09° off axis according to the Code V ray-tracing model of the WHAM optical system. The off-axis angles are measured at the etalons.

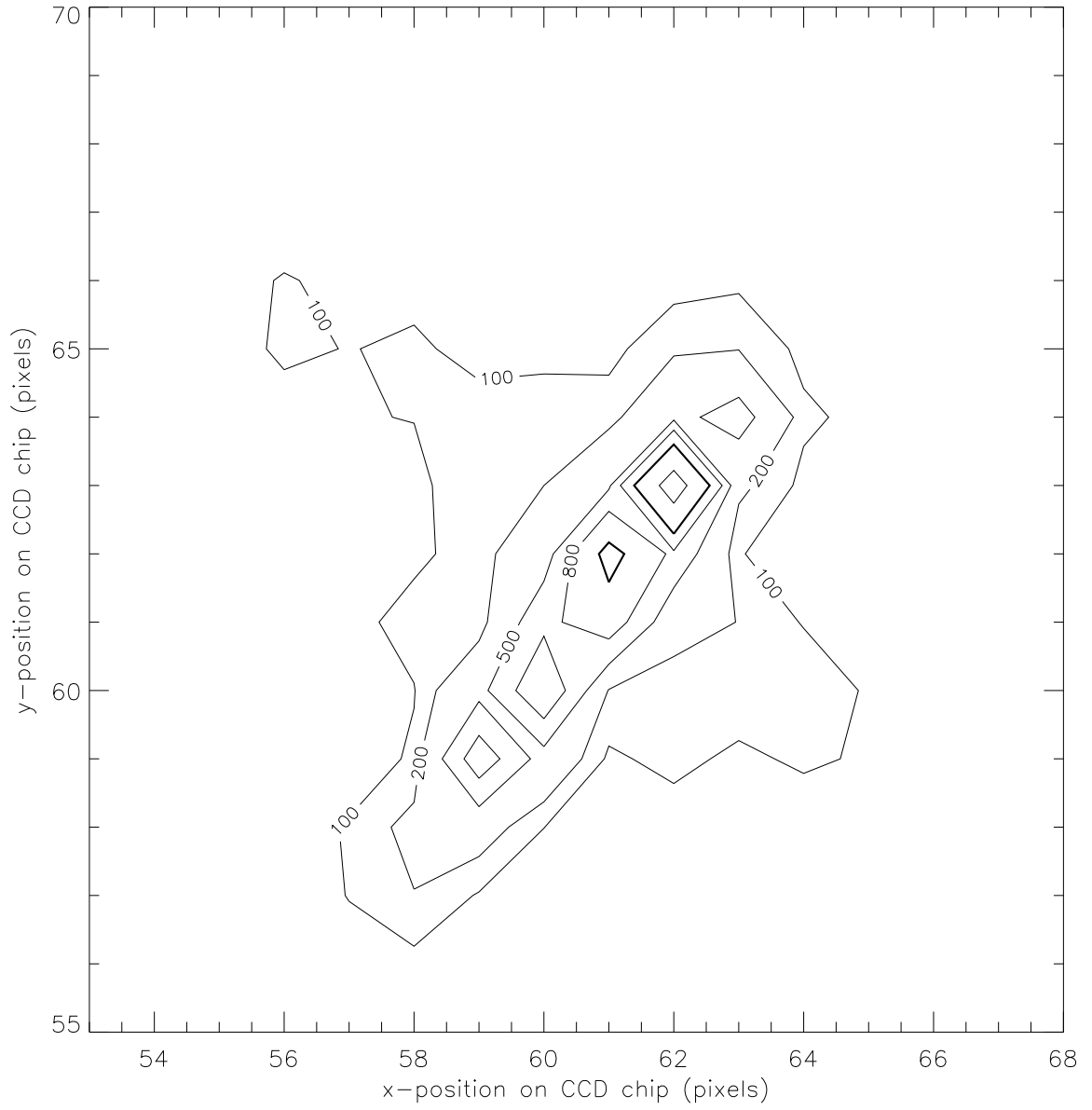


Fig. 3.15.— Contour plot of a star image.

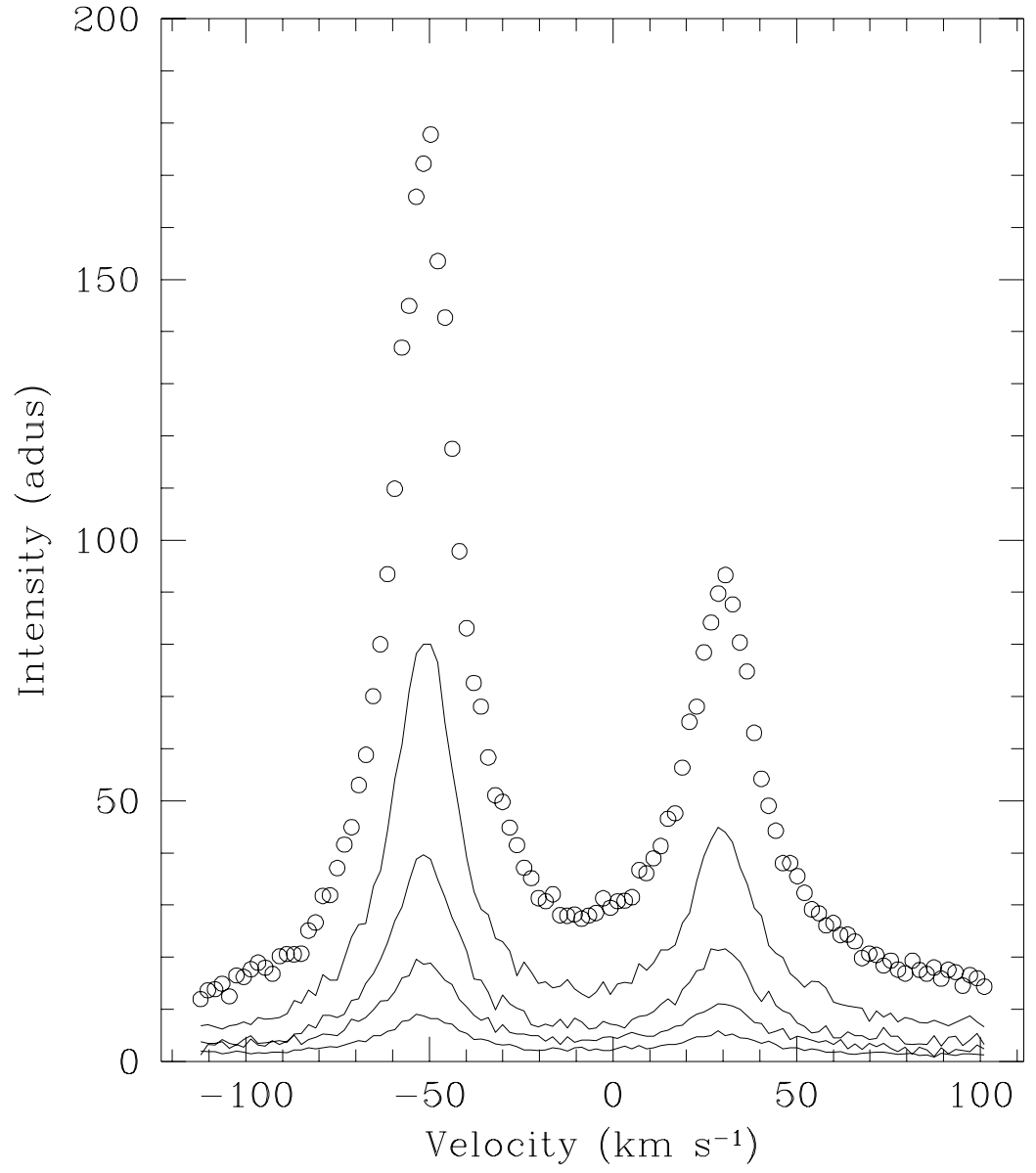


Fig. 3.16.— Spectra of $H\alpha - D\alpha$ lamp at 120s (circles), 60s, 30s, 15s, and 7.5s exposure times.

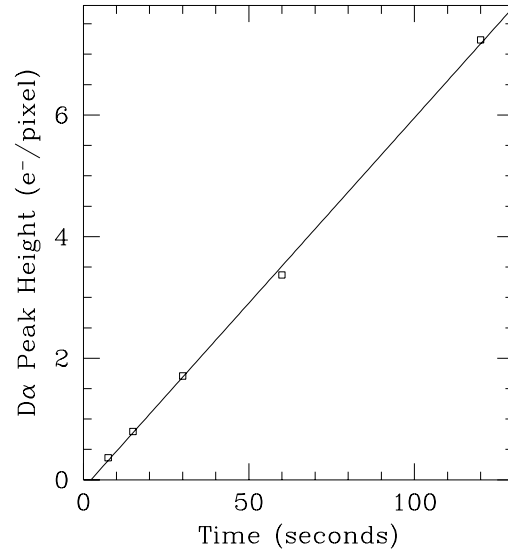


Fig. 3.17.— Peak Height of D α Gaussian vs. exposure time. The height is expressed in units of e^- per 1×1 pixel. The curve shows a linear fit to data.

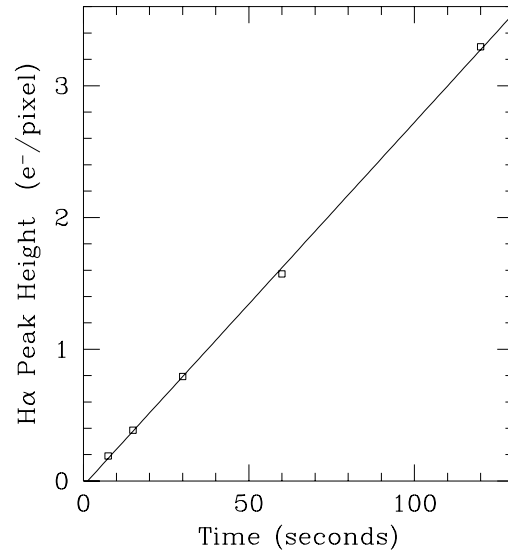


Fig. 3.18.— Peak Height of H α Gaussian vs. exposure time. The height is expressed in units of e^- per 1×1 pixel. The curve shows a linear fit to data.

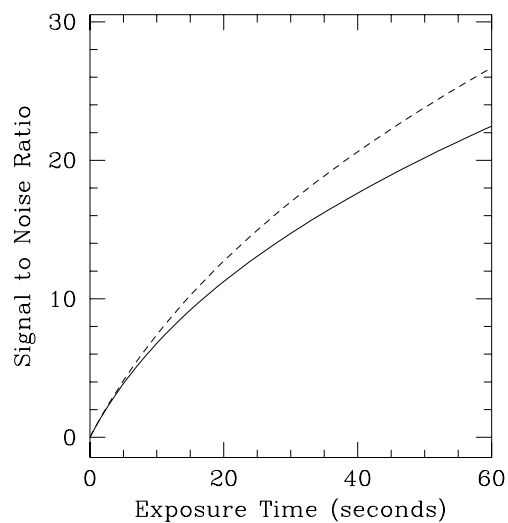


Fig. 3.19.— Expected S/N per data point versus exposure time for a 1 R line. The solid curve corresponds to the Madison night sky brightness and dashed curve is based on the estimated night sky brightness at Kitt Peak.

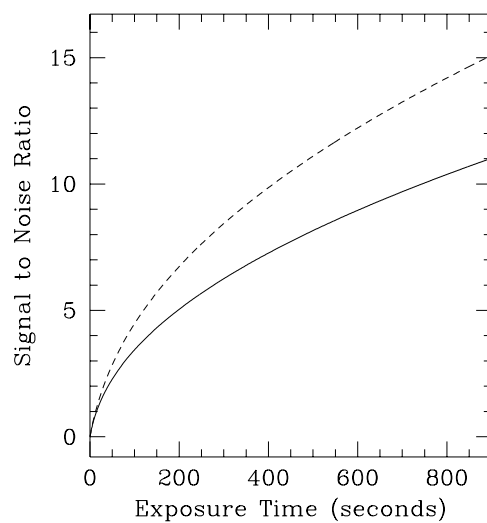


Fig. 3.20.— Expected S/N per data point versus exposure time for a 0.1 R line. The solid curve corresponds to the Madison night sky brightness and dashed curve is based on the estimated night sky brightness at Kitt Peak.

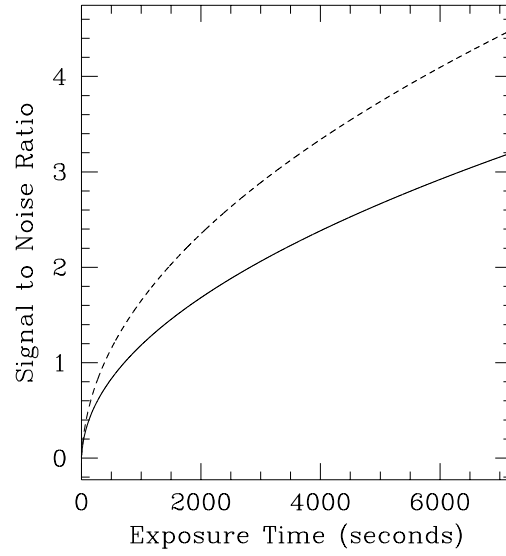


Fig. 3.21.— Expected S/N per data point versus exposure time for a 0.01 R line. The solid curve corresponds to the Madison night sky brightness and dashed curve is based on the estimated night sky brightness at Kitt Peak.

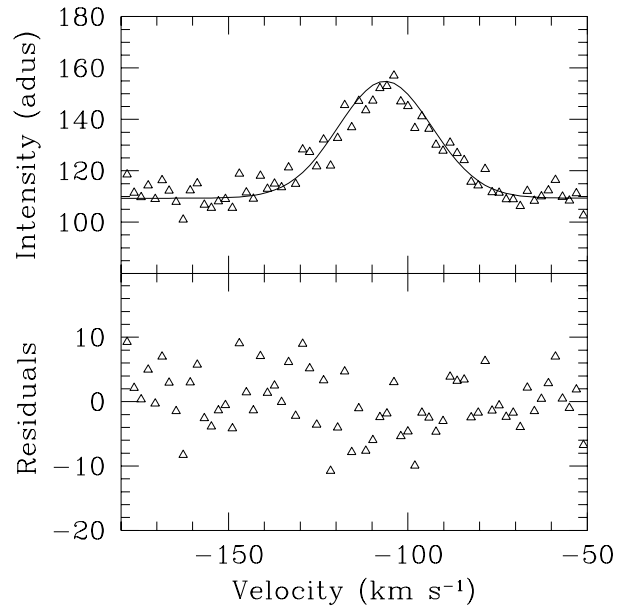


Fig. 3.22.— $H\alpha$ spectrum of HVC M I with residuals.

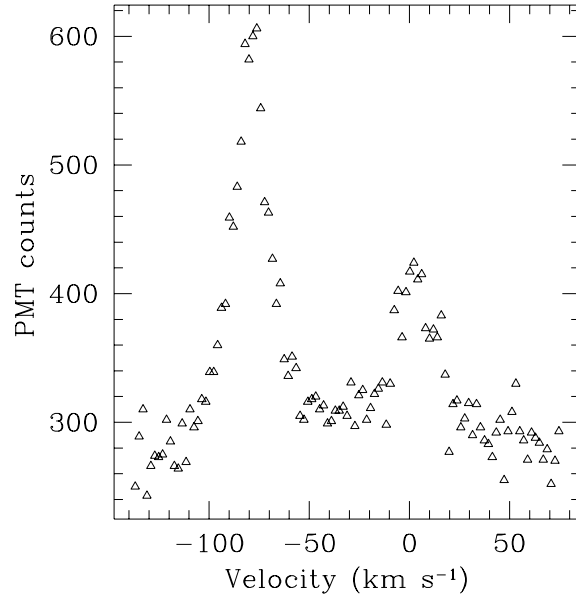


Fig. 3.23.— Old PMT based instrument spectrum from a 1090 second scan of the $\simeq 1$ R H α -D α diffuse light source.

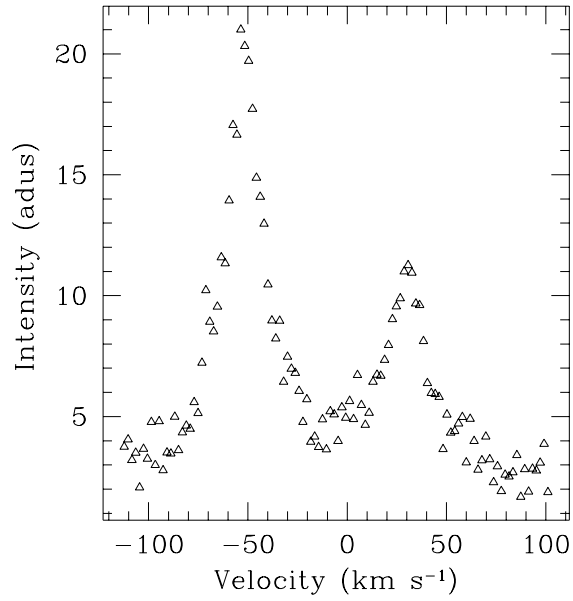


Fig. 3.24.— WHAM instrument spectrum from a 15 second exposure of the $\simeq 1$ R H α -D α diffuse light source.

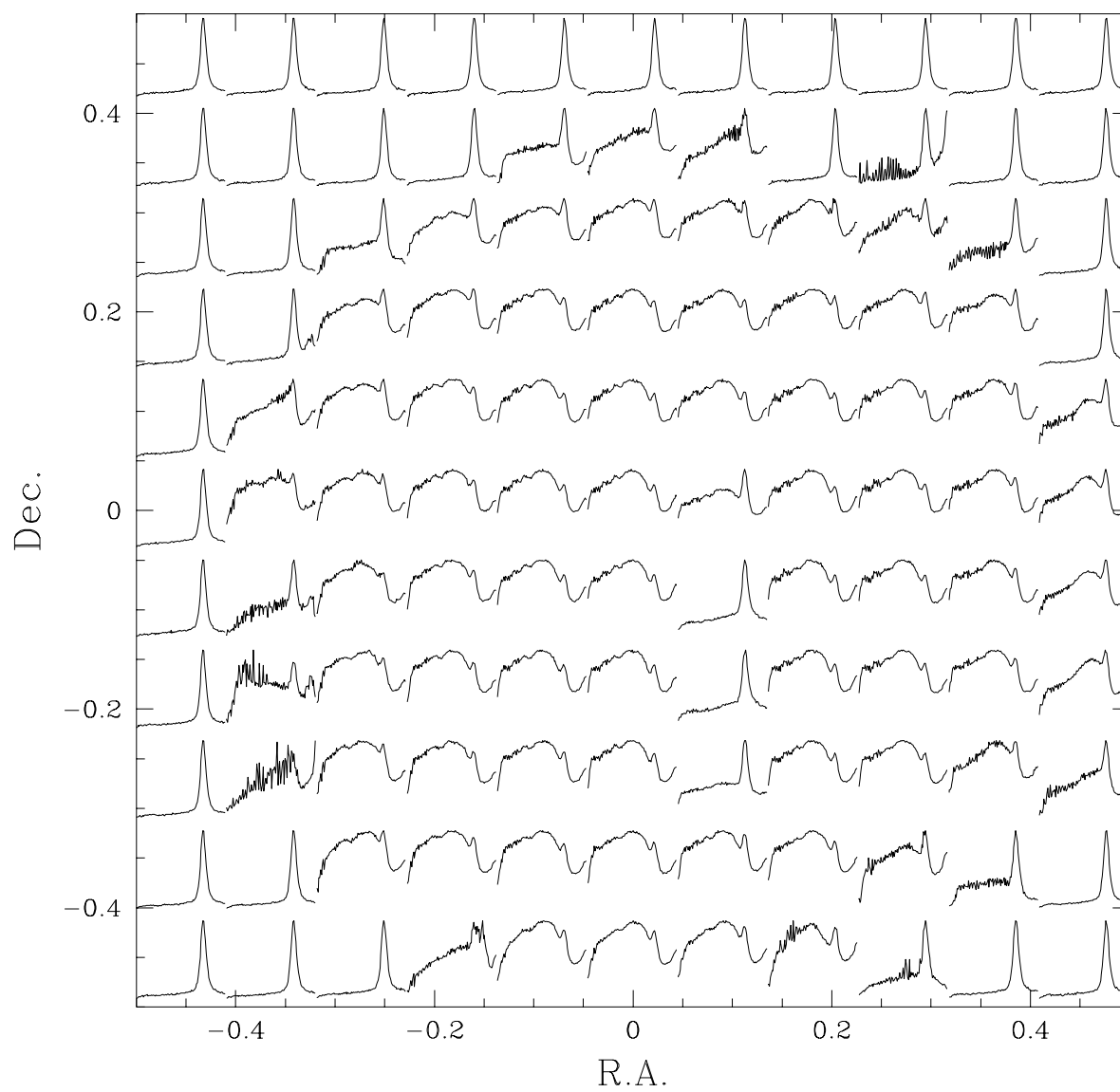


Fig. 3.25.— Spectra of α Boo near $H\alpha$ with the star placed at a grid of locations with respect to the beam center.

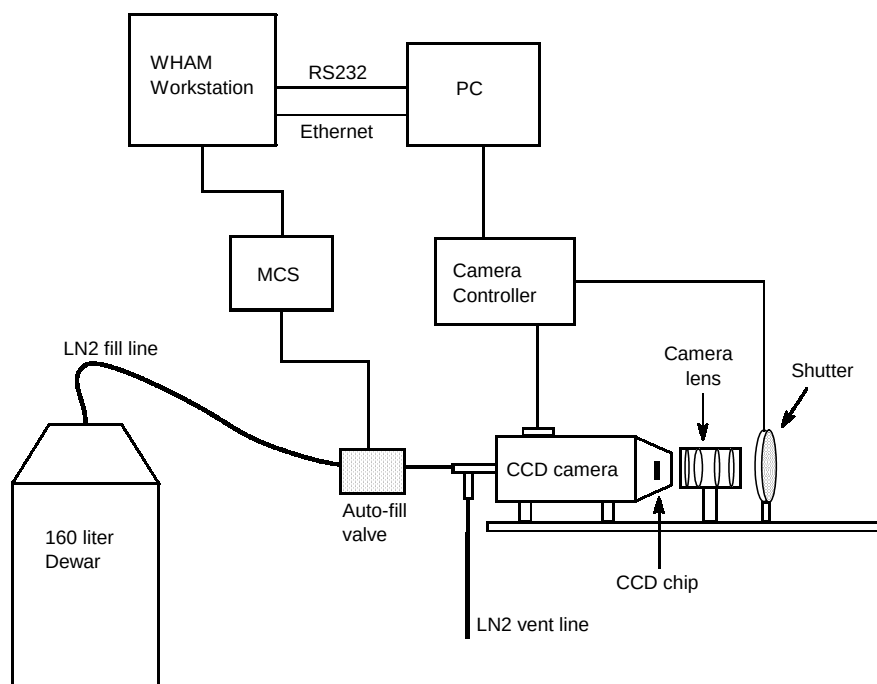


Fig. 3.26.— The WHAM CCD camera system.

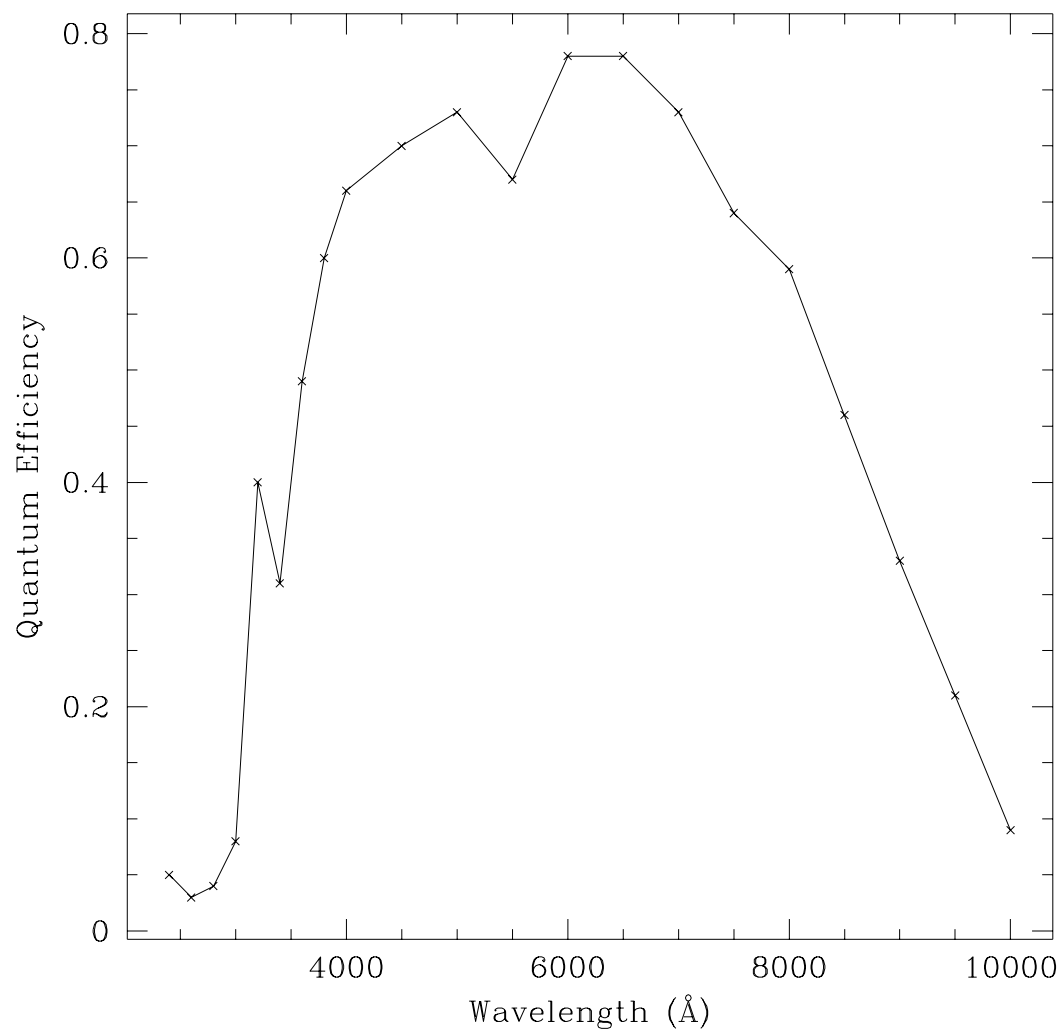


Fig. 3.27.— Quantum efficiency vs. wavelength for the Tektronix 1024AB2 chip.

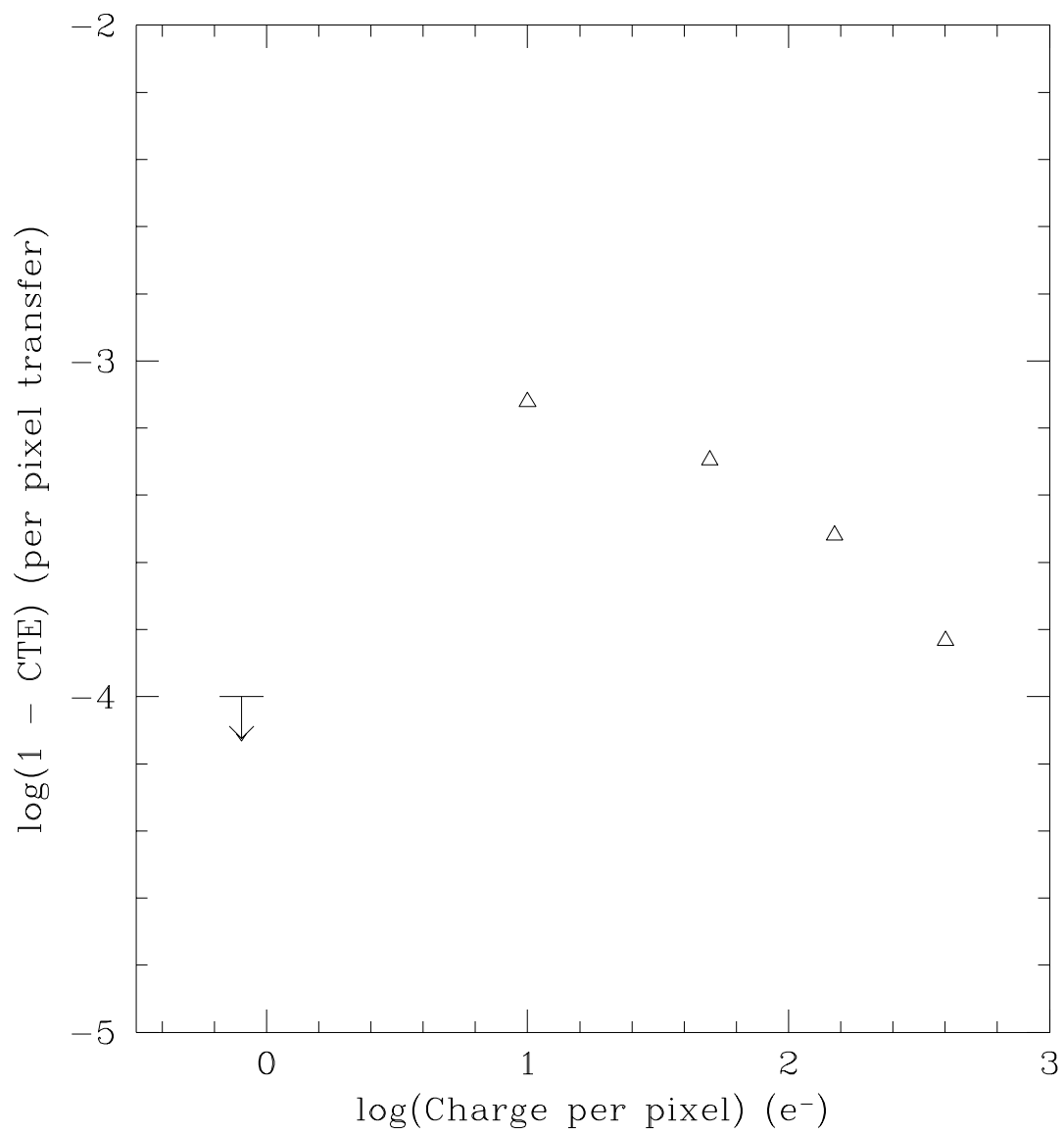


Fig. 3.28.— $(1 - \text{CTE})$ versus charge level per pixel.

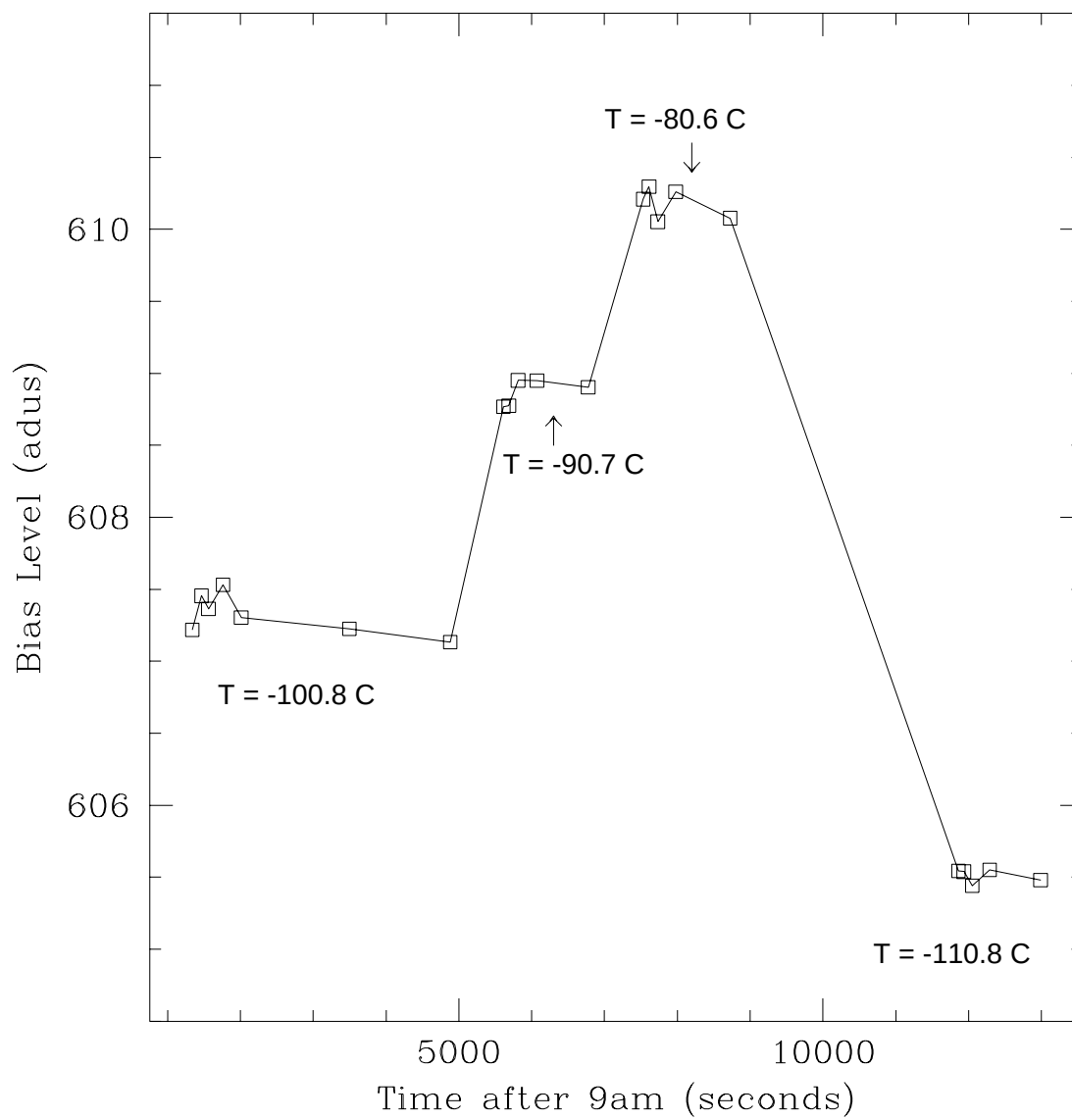


Fig. 3.29.— Bias level versus time at various temperatures.

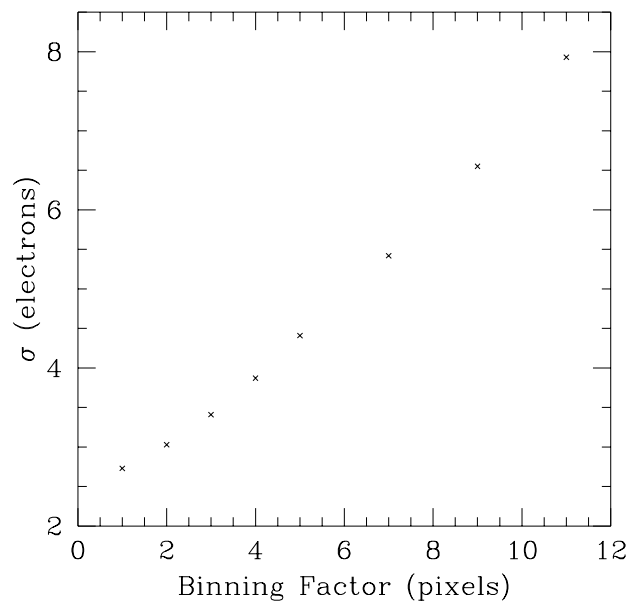


Fig. 3.30.— Bias noise vs. binning factor. The σ obtained in a bias image as a function of the linear binning factor (e.g. a binning factor of b implies the bin size was $b \times b$).

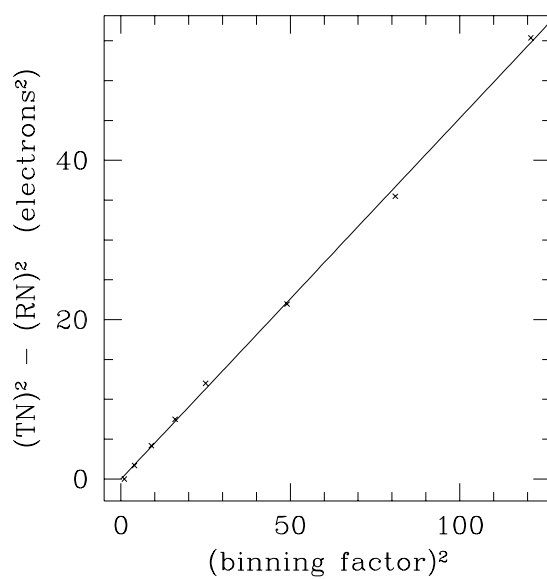


Fig. 3.31.— $(\text{Total Noise})^2 - (\text{Read Noise})^2$ vs. $(\text{Binning Factor})^2$, square bins only.

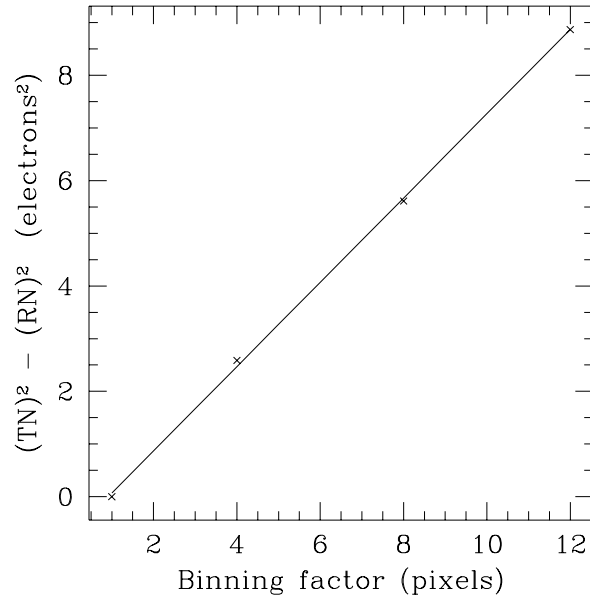


Fig. 3.32.— $(\text{Total Noise})^2 - (\text{Read Noise})^2$ vs. serial binning factor where the parallel binning factor is 1 (e.g. binned pixels are $b \times 1$).

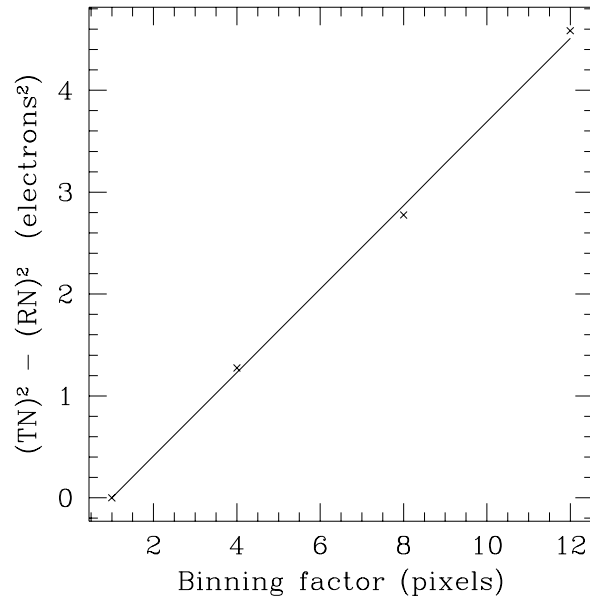


Fig. 3.33.— $(\text{Total Noise})^2 - (\text{Read Noise})^2$ vs. parallel binning factor where the serial binning factor is 1 (e.g. binned pixels are $1 \times b$).

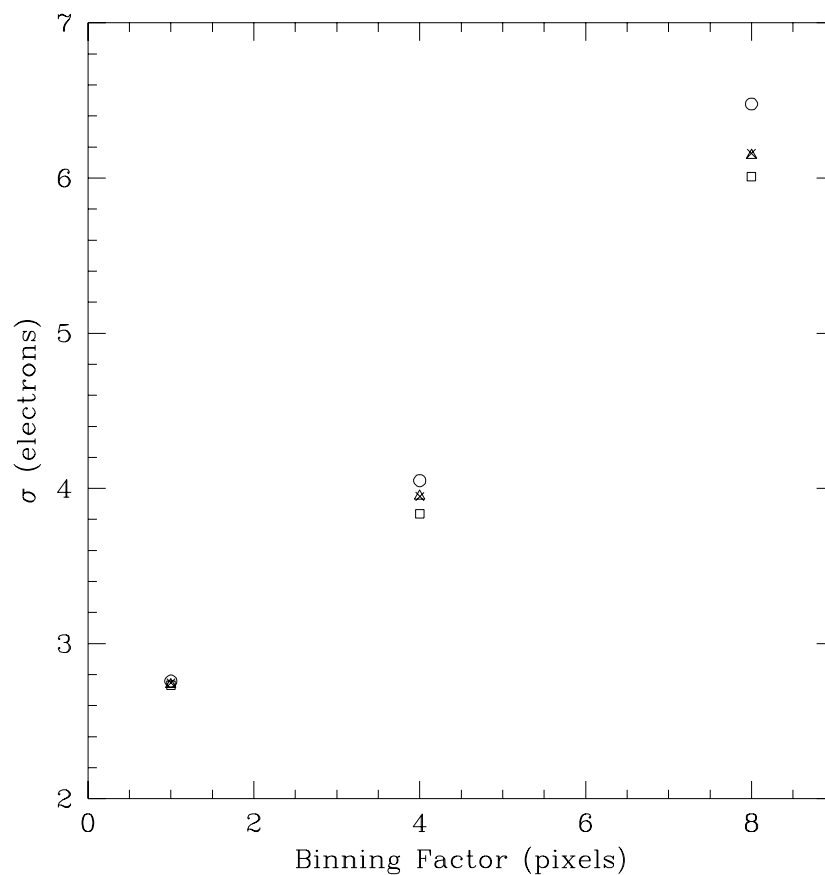


Fig. 3.34.— σ vs. binning factor for 4 subregions of the chip. triangles = lower left quadrant (location of amplifier), squares = lower right quadrant, X = upper right quadrant, and circle = upper left quadrant. Parallel transfers are top to bottom, and serial transfers are right to left.

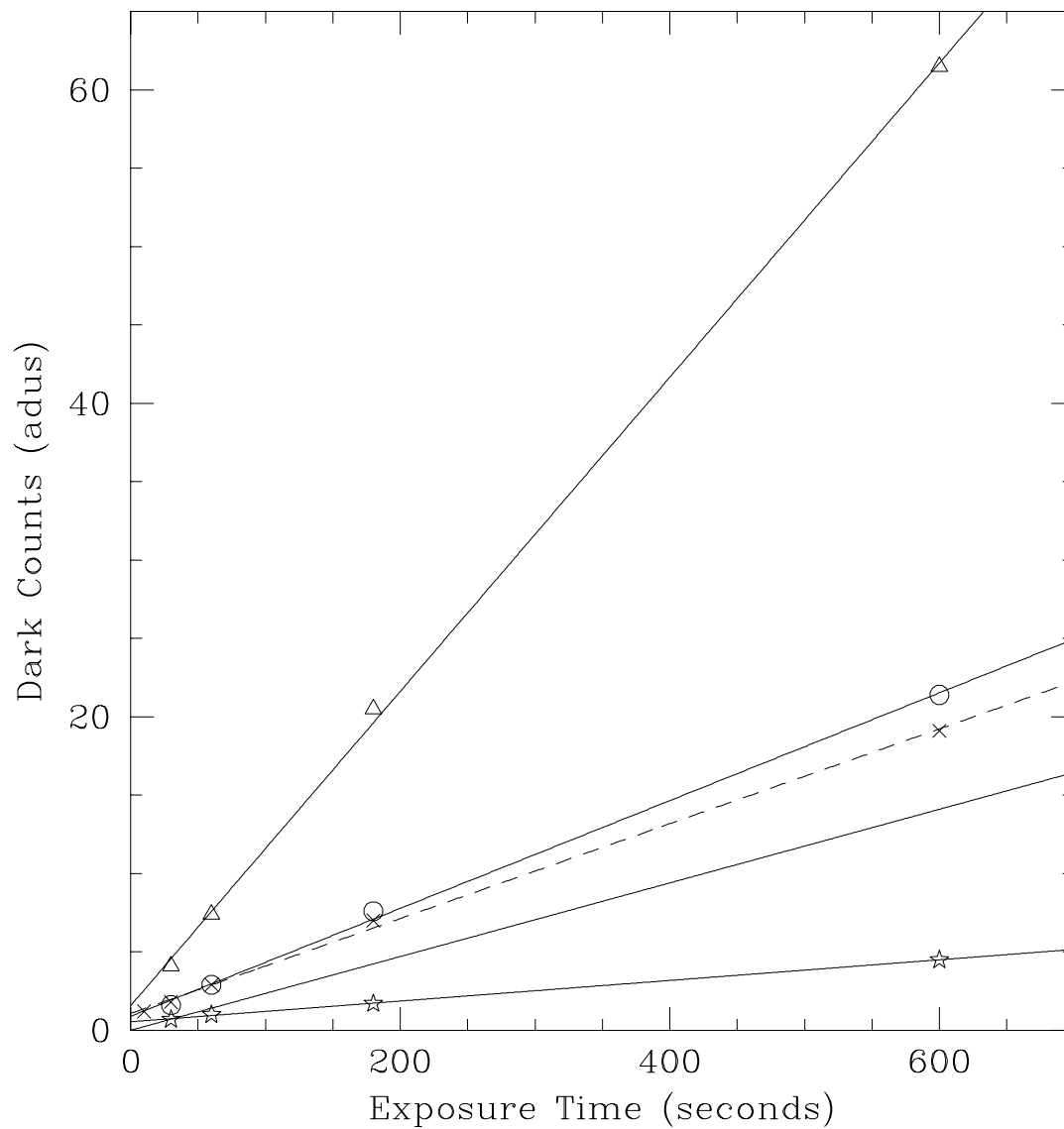


Fig. 3.35.— Dark counts versus time. The solid lines are for $T = -80.6^{\circ}\text{C}$, $T = -90.7^{\circ}\text{C}$, $T = -100.8^{\circ}\text{C}$, and $T = -110.8^{\circ}\text{C}$ (top to bottom). The dashed line shows a fit to the first five data points gathered at $T = -100.8^{\circ}\text{C}$ (see Figure 3.36 and the discussion in Section 3.8.7).

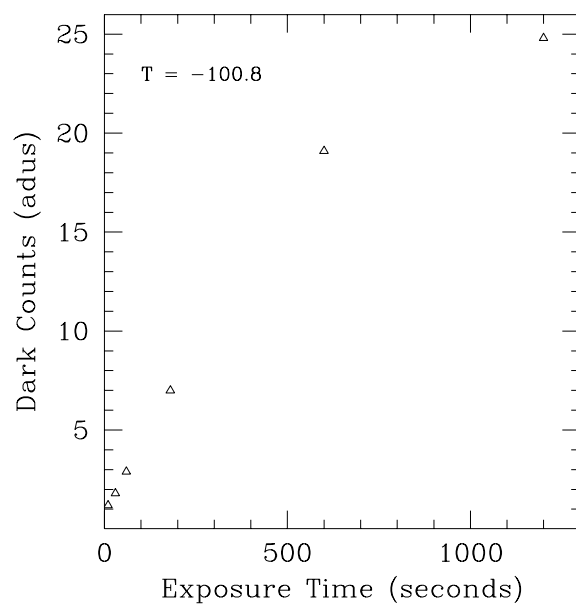


Fig. 3.36.— Dark counts versus time for $T = -100.8^{\circ}\text{C}$.

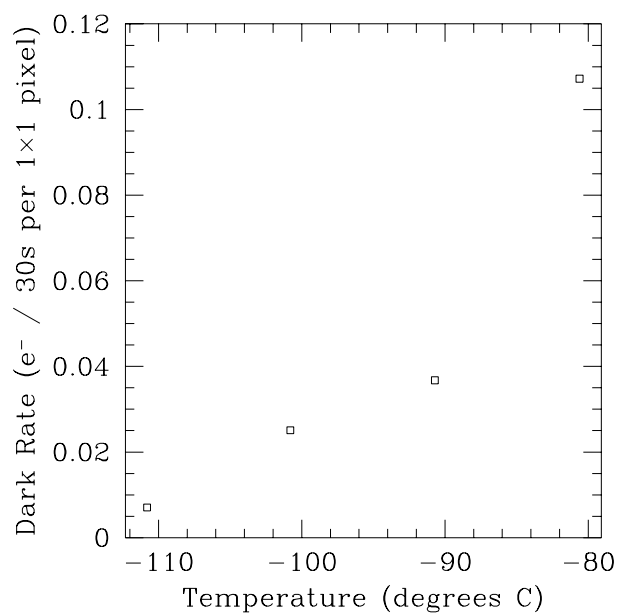


Fig. 3.37.— Dark rate versus temperature.

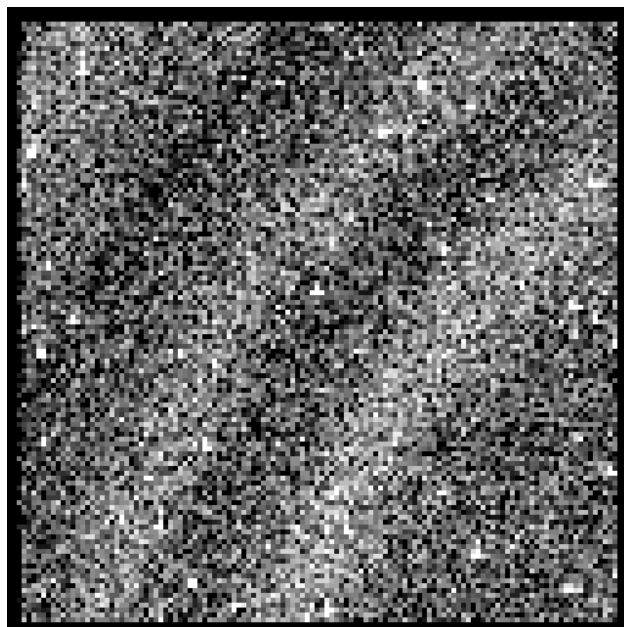


Fig. 3.38.— 600 second dark image at $T = -80.6^{\circ}\text{C}$.

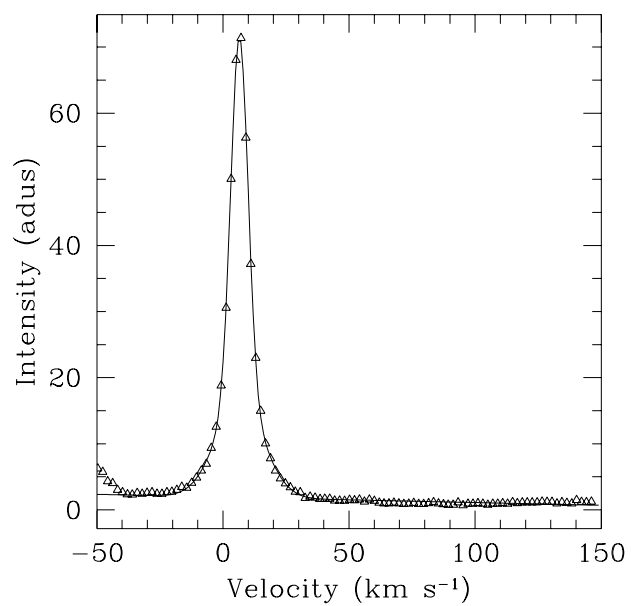


Fig. 3.39.— Spectrum of $\lambda 5753$ thorium line near optimal focus.

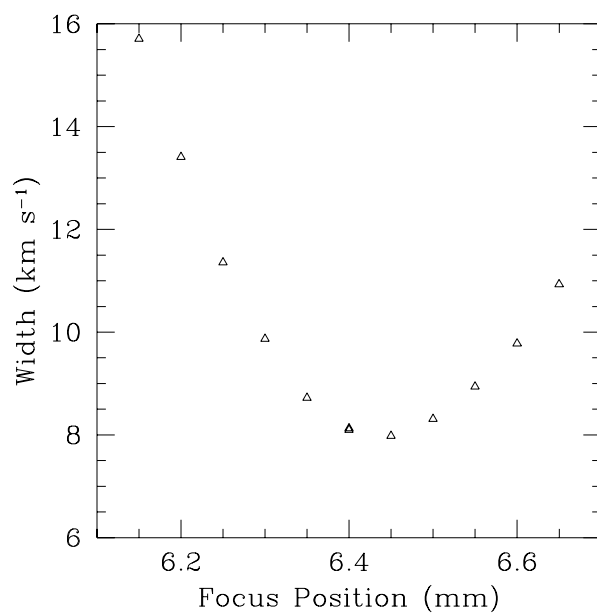


Fig. 3.40.— Width of thorium $\lambda 5753$ line vs. focus position.

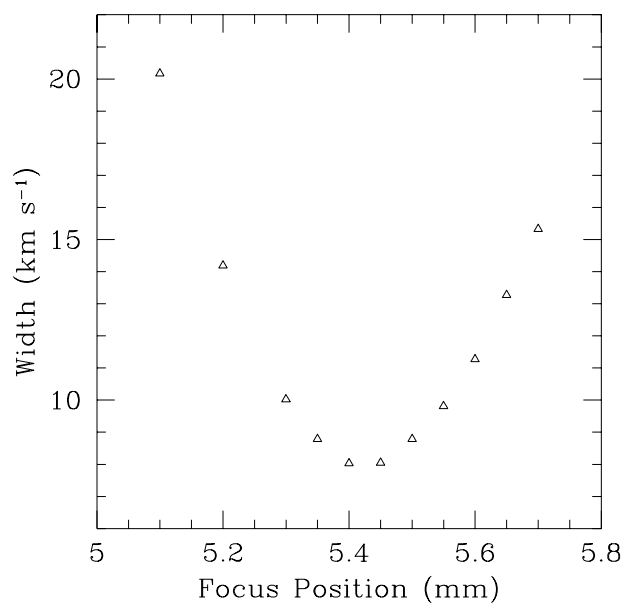


Fig. 3.41.— Width of thorium $\lambda 4863$ line vs. focus position.

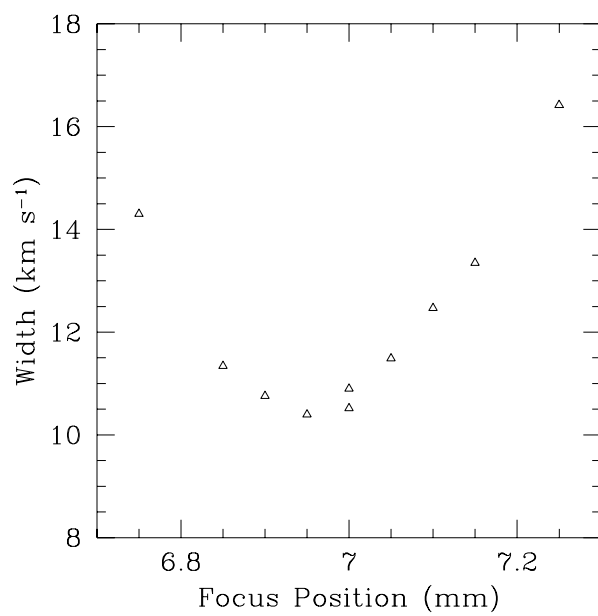


Fig. 3.42.— Width of thorium $\lambda 6554$ line vs. focus position.

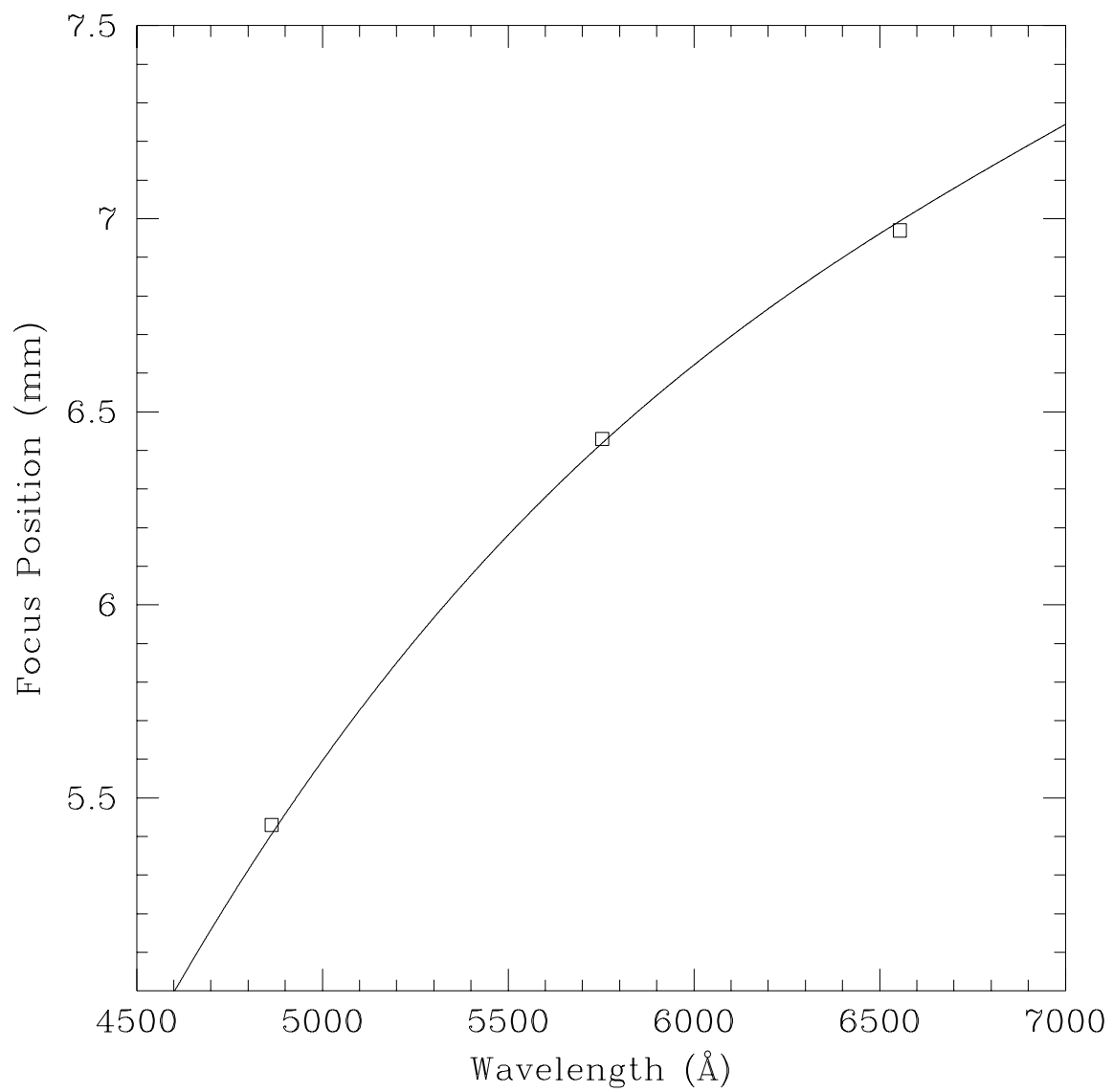


Fig. 3.43.— Best focus position vs. wavelength. Squares: measured, curve: model using Code V.

Chapter 4

WHAM Observations of Emission Lines from HVCs in the M Complex

4.1 Introduction

High velocity clouds (HVCs) were discovered over 30 years ago through the 21 cm line of neutral hydrogen (Muller, Oort, and Raimond 1963). While they have been studied extensively (and almost exclusively) in the 21 cm line since then, the origin and nature of HVCs is still largely unknown. They are observationally defined as gas moving at velocities not explainable by cylindrical galactic rotation (a rule of thumb is $|v_{lsr}| \geq 80 \text{ km s}^{-1}$). In the northern hemisphere, the vast majority of the HVCs have negative velocities, that is, they are falling toward the Galactic plane. In addition, there exists gas at intermediate velocities (IVCs), also primarily falling toward the plane in the northern hemisphere, which may have a connection with the HVCs (see Kuntz & Danly 1996). Many theories have been proposed to explain the HVCs. One model suggests that they are primordial intergalactic hydrogen clouds

captured by the gravitational field of the Galaxy (Oort 1966, Oort 1970, Wakker 1990). Another suggests that they are the cooled condensate of material vented into the halo as a result of super-bubble breakout (the galactic fountain model, see Shapiro & Field 1976, Bregman 1980). Another proposal (Murai & Fujimoto 1980, Giovanelli 1981) is that the gas is tidally stripped from the Magellanic clouds as they orbit the galaxy (there is convincing evidence that this is the case for a southern complex of high velocity gas known of as the Magellanic Stream, but it is harder to associate the northern HVCs with this origin.)

A sensitive all-sky survey of HVCs now exists (Hulsbosch & Wakker 1988), and the HVCs have been catalogued and organized into regions called complexes. In this chapter, we report on our studies of optical emission lines from Complex M. This region is especially interesting because of its high galactic latitude ($50^\circ \leq b \leq 70^\circ$). At this latitude, galactic rotation cannot be a significant factor in the velocity, and the radial velocity is a fairly good measure of motion perpendicular to the galactic plane. Further, dust obscuration is less significant than at lower latitudes.

One of the biggest problems limiting progress in understanding the HVCs is the lack of knowledge of the distances to the emitting gas. The usual method of determining distances to neutral hydrogen is to use the kinematic distance derived from a knowledge of the differential rotation of the galaxy. This method fails for high velocity clouds since by definition they do not fit the usual pattern of Galactic rotation. The only reliable method of determining distances to HVCs is absorption line spectroscopy toward bright objects of known distance in the direction of the cloud. If an absorption component is detected at the known velocity of the cloud under study, then one can infer that the cloud must be closer than the object. If the component

does not appear, this suggests that the cloud is more distant than the light source; however, in this case one must consider the possibility that the line of sight has passed through a break in the potentially patchy cloud. By choosing bright objects at various distances, it is possible to constrain the distance to the cloud.

As an example, consider the measurements of Danly et al. (1993) toward the M complex. They measured absorption spectra toward two stars separated by $27'$ on the sky in the direction of the M complex using the IUE satellite. The more distant object, BD +38°2182, is at a height z above the galactic plane of 4.4 kpc and shows absorption features in saturated resonance lines of C II, S II, and O I at a velocity that matches the the 21 cm emission of the HVC seen in the same direction. The other object, HD 93521, is at $z = 1.5$ kpc and shows no high velocity absorption. The authors argue that the nearness of the two objects on the sky and the smoothness of the 21 cm emission in the region make it unlikely that the cloud is in front of the HD 93521 sightline. Hence, they conclude that the gas along their sightline is between $1.5 \leq z \leq 4.4$ kpc. They also generalize this result to the entire M complex. If these distances are correct, the M complex clouds are excellent “test particles” for probing the conditions in the Galactic halo.

An analysis of Rosat observations of soft X-rays combined with 21 cm data from the new Leiden/Dwingeloo survey toward Complex M by Herbstmeier et al. (1995) suggests that there is a brightening of soft X-rays at several cloud locations, perhaps due to collisions of the clouds with the ambient medium in the halo. The most striking such enhancement occurs on the edge of cloud M I, where the H I contours are also seen to be crowded. It should be noted, however, that considerable modeling is necessary to separate out the soft X-rays possibly associated with the M complex region from

sources of soft X-rays at other distances in the same direction.

While considerable effort has been expended looking for emission lines from HVCs in the past, there have only been a few detections. A search for $H\alpha$ by Reynolds (1987) in 6 directions toward HVCs, including two directions toward the M complex, yielded a negative result. The upper limits placed on the intensity by this study ranged from 0.2 R to 0.6 R. Kuttyrev and Reynolds (1989) obtained a 4σ detection of a very high velocity cloud ($v_{lsr} = -300 \text{ km s}^{-1}$) in Cetus. The $H\alpha$ brightness in this direction was measured to be $I_\alpha = 0.08 \text{ R}$. Münch & Pitz (1989) marginally detected high velocity $H\alpha$ emission from the M II cloud with $I_\alpha = 0.15 \text{ R}$ (one of our observation directions is very near theirs, and I will compare the results below). Songaila et al. (1989) claim a detection of high velocity gas in Complex C. The intensity they measured was 0.03 R; however, Reynolds (1996) has since observed the same direction and found the emission to be significantly brighter than the Songaila & Cowie measurement. Very recently, Weiner & Williams (1996) have detected $H\alpha$ emission from a number of directions toward the Magellanic Stream. They find that $H\alpha$ emission enhancements (typically $\simeq 0.2 \text{ R}$) are associated with gradients in the H I distribution, which they interpret as evidence that the emission is produced by ram pressure heating of the cloud as it passes through a low density ambient medium. In general, the previous observations of $H\alpha$ from HVCs required long integration times and, with the exception of the Magellanic Stream detections, produced null or marginal results.

The recent development of CCD detectors with very low read noise and dark current provide us with the capability to explore optical emission lines with unprecedented sensitivity. This capability opens a new window to the HVCs. Much can be

learned about both the HVCs and the Galactic halo environment from a study of optical emission lines. For example, the brightness of the $H\alpha$ line gives information on the amount of ionized hydrogen that is associated with the neutral gas. If the ionizations are due to photoionization, then the $H\alpha$ intensity is a direct measure of the Lyman continuum flux incident upon the cloud. If the ionizations are due to collisional processes induced by the rapid motion of the cloud (i.e. shocks), then the intensity is a measure of the ambient density of material through which the cloud travels. It may be possible to distinguish between these two ionization mechanisms by measuring other emission lines such as [S II] $\lambda 6716$, [N II] $\lambda 6584$, and [O III] $\lambda 5007$. It is also possible that the pattern of $H\alpha$ intensity may shed some light on the relationship between the neutral and ionized gas and perhaps features in the ISM at lower velocity.

The new WHAM spectrometer is well suited to the study of emission lines from HVCs for several reasons: 1) the Fabry-Perot spectrometer is ideal for detecting spatially extended emission from faint sources at high resolving power due to the much higher throughput available at a given spectral resolution compared to grating spectrometers, 2) the 15 cm etalons used by the WHAM spectrometer are the largest available, 3) the 1° beam is well matched to the large angular sizes characteristic of many of the HVCs, and 4) the 12 km s^{-1} resolution is well matched to the 20–25 km s^{-1} (FWHM) line widths (21 cm) typical of HVCs. As a result, the WHAM spectrometer allows us to obtain clear detections of emissions with intensity of order 0.1 R in 30 minutes integration time.

We have used the recently completed WHAM spectrometer to detect $H\alpha$ in 7 directions toward HVCs in the M complex. We have also detected the [S II] $\lambda 6716$ line from the M I cloud. In this chapter I present the results of our studies of HVCs

in Complex M and briefly discuss the implications of these measurements.

4.2 Observations

4.2.1 The $H\alpha$ Observations

We obtained $H\alpha$ observations in 15 directions in the vicinity of the high-velocity clouds M I and M II. Table 4.1 lists the coordinates of each look direction along with the total exposure time and the date(s) of observation. Figure 4.1 shows our look directions on a map of the 21 cm contours provided by Bart Wakker. The 21 cm data, from Hulsbosch and Wakker (1988), measure the amount of neutral hydrogen at high velocity ($v_{lsr} \leq -80 \text{ km s}^{-1}$). The contour levels for the 21 cm brightness temperature data are 0.04, 0.3, 0.6, 1.5, 2.0, and 3.0 K, corresponding to $N_{HI} \simeq 1.5 \times 10^{18}$, 1.1×10^{19} , 2.3×10^{19} , 5.7×10^{19} , 7.7×10^{19} , and $1.1 \times 10^{20} \text{ cm}^{-2}$, respectively (assuming a line width of 20 km s^{-1}). The $H\alpha$ study was conceived as a set of 8 “on–off” pairs. The directions 1a, 2a, 5a, 6a, 7a, and 8a are “on” the M I cloud as defined by the 21 cm maps, and the associated “off” directions, which are 4a, 3a, 9a, 12a, 11a, and 10a, were chosen based on their lack of detected H I gas at high velocity. Note that direction 11a, located at $l = 140.0^\circ$, $b = 80.0^\circ$, does not appear on this map. A more detailed look at the location of these beams is shown in Figure 4.2. For the M II cloud, the “on” directions are 1b and 2b, and the “off” direction is 3b (see Figure 4.1).

In order to isolate the emission due to the HVCs under study, we use an “on” minus “off” method. This requires an observation “on” the direction of interest, immediately followed by an observation of a nearby direction that is “off” the high velocity cloud. This very powerful technique allows for the subtraction of the sky

spectrum and cancelation of many systematic uncertainties. Since the intermediate velocity gas is fairly smooth in the Complex M region (at least in the 21 cm maps), this method also serves to subtract off the IVC component, allowing for the isolation of the HVC component. Figure 4.3 shows an example of the method. The upper panel depicts the spectrum of an “on” direction (derived from a 900 second exposure of the 1a direction) as a solid line and the spectrum of an “off” direction (derived from a 900 second exposure of the 4a direction taken immediately after the 1a exposure) as a dashed line. Both spectra show the relatively bright geocorona on the right and very little else. The geocoronal emission line is placed near the red edge of the bandpass for use in the velocity calibration. These spectra have not been flat fielded, and as a result they both decrease in brightness at the extreme left edge of the spectrum (corresponding to the edge of the aperture) because of a slight misalignment of the optical system which leads to underillumination of the edge of the aperture defined by the filter (now corrected). The lower panel shows the result of subtracting the “off” spectrum from the “on” spectrum after flat fielding. Here one can clearly see an emission component near -100 km s^{-1} due to the M I cloud. On the right, the difference spectrum takes a steep dive due to a combination of the fact that the geocoronal line is brighter in the “off” direction and that the 4a direction has more emission from intermediate velocity gas (see below). In general, problems with accurately correcting for the roll off at the edge of the aperture make it difficult to measure extremely faint emission lines at velocities between -200 km s^{-1} and -175 km s^{-1} . Problems with accurately removing the geocorona make it difficult to extract information at velocities greater than about -40 km s^{-1} . This leaves a velocity interval of 130 km s^{-1} within which we can make good measurements of very faint emissions from high velocity gas.

We discovered several emission lines in the night sky spectrum that are probably extremely weak, not previously detected OH lines. Figure 4.4 shows the average of two spectra from the February 24 observations. On this night, the sky lines were the brightest, although the same lines are present to some degree in all of the spectra. The evidence for the lines originating in the atmosphere and not the Galaxy is 1) The lines are present in all of the spectra (both “on” and “off”), 2) their brightness seems to depend more on the night than the look direction, 3) the location of the lines is fixed with respect to the geocorona and not fixed with respect to the LSR, and 4) the lines appear to be very narrow, as would be expected for atmospheric lines but not Galactic emission. The night sky features were fitted with Gaussian components at locations shown by the arrows in Figure 4.4. A terrestrial spectrum was then subtracted from each spectrum as part of the data reduction process. This spectrum was formed using the spectrum shown in Figure 4.4 as a template. The locations and relative strengths of the sky lines were held fixed and a single parameter corresponding to the overall brightness of the sky lines was adjusted by fitting the brightest line (the second from the left) in each spectrum. The analysis of the “on–off” pairs was done with and without this night sky line subtraction to insure that the process was not introducing spurious emission features. The results for the “on–off” spectra were very similar in the two cases, as is expected since the sky line intensity changes little on hour time scales. The data analysis for the night sky line subtraction was done by Nikki Hausen under the guidance of Ron Reynolds.

4.2.2 The S II Observations

For directions 1a, 4a, 6a, and 12a we also obtained spectra in the [S II] $\lambda 6716$ line. On April 10, 1996 two 900 second exposures were obtained for each of these directions.

4.3 Results

In this section I present the results of our “on–off” pair observations for the $H\alpha$ and [S II] lines. I also describe our preliminary attempts at measuring the absolute level of emission as a function of velocity by using the Lockman window direction (10a) spectrum as the zero level.

4.3.1 The $H\alpha$ “on–off” Spectra

The spectrum obtained by averaging the 5 “on–off” pairs measured for the 1a–4a direction is shown in Figure 4.5. An emission line is clearly detected at a velocity with respect to the local standard of rest (LSR) $v_{lsr} = -106 \text{ km s}^{-1}$. For $v_{lsr} \leq -170 \text{ km s}^{-1}$ the spectrum rises due to our current inability to accurately correct for the roll-off caused by vignetting. For $v_{lsr} \geq -50 \text{ km s}^{-1}$ the spectrum decreases drastically due to differences in the intermediate velocity gas intensity and the geocoronal intensity between the “on” and the “off” directions (i.e., the $H\alpha$ emission associated with the lower velocity gas is brighter in the “off” direction). A Gaussian fit to the emission line, convolved with the instrumental response function, is shown as the solid curve in the figure. The extent of the curve shows the velocity range included in the fit. The area under this curve leads to an intensity for this emission line of 0.078 R. The full width at half maximum (FWHM) of the fitted Gaussian is 27 km s^{-1} . It must be

emphasized that this is the difference between the emission seen in the 1a direction and the 4a direction and that the actual emission spectrum for the 1a direction alone would be quite different (see below). The intermediate velocity gas (at least the neutral portion) is a fairly uniform sheet in this part of the sky, and so subtracting the 4a direction allows us to extract just the emission due to the the M I cloud. The arrow in the figure shows the location of the 21 cm emission line from neutral hydrogen interpolated from observations near the 1a direction (the Hulsbosch and Wakker 21 cm data contains 0.5° beams on a 1° grid.) The velocities of the emission from the neutral and ionized gas agree to within the experimental uncertainties in this case. Table 4.2 summarizes the results from all of the “on-off” pair analysis.

In order to get a handle on the systematic uncertainties, I fitted each of the 5 “on-off” pairs for this direction separately and then compared the scatter obtained in the mean, width, and area parameters of these fits to the uncertainties reported by the fitting program for these parameters. This exercise is interesting because each of the 5 independent measurements were gathered under slightly different conditions. For instance, the night sky spectrum varies from night to night. While the “on-off” technique is designed to subtract out the sky spectrum, this will not work perfectly because of the smaller time scale variation. Therefore, on a night when the atmospheric lines are bright they might not subtract out as well as on a night when the lines are faint. Note that the brightest terrestrial line, on the night when the sky line contamination was the worst (2/24/96), is slightly brighter than the M I cloud emission and that the sky line that is in the same region of the spectrum as the HVC emissions is about $1/2$ as bright as the brightest sky line. On all other nights, the brightest line is $\lesssim 1/3$ of the M I cloud intensity. Another factor is that the location of the emission

from the M I cloud on our spectrum changed by approximately 20 km s^{-1} between the first observation and the last. This is due to the fact that we always arranged to have the geocorona on the right edge of the spectrum, but the motion along the sightline with respect to the local standard of rest changed due to the changing projection of the earth's orbital motion. Hence the location of the ring due to the HVC emission line falls on slightly different parts of the CCD chip for the different “on-off” pairs, and therefore there could be systematic effects which varied with location within the bandpass. All of these effects will tend to make the scatter among the different “on-off” pairs larger than the uncertainties derived from a given pair.

In the case where the 5 “on-off” pairs were fitted separately, the average intensity was 0.0838 R with a scatter of $\sigma_I = 0.016 \text{ R}$. The average value of the uncertainty reported by the fitting program for the intensity is 0.014 R. The average of the widths determined from the 5 fits is 28.5 km s^{-1} and $\sigma_W = 6.7 \text{ km s}^{-1}$. The average of the uncertainties in the width reported by the fitting program is 5.4 km s^{-1} . For the mean, the average value was 106.1 with $\sigma_M = 1.0 \text{ km s}^{-1}$. This is a strong confirmation that the emission line is of galactic origin, since the earth's motion along the line of sight changed by around 20 km s^{-1} . The fitting program mean uncertainty is 1.7 km s^{-1} . These values are in reasonable agreement, implying that the uncertainties calculated by the fitting program can be used to get an estimate of the uncertainties involved in measuring the parameters of the emission lines. Dividing the σ values determined above from the scatter in the different “on-off” pairs by $\sqrt{5}$ gives an estimate of the uncertainty in the averages of the values. These numbers are $\delta I = 0.007 \text{ R}$, $\delta W = 3.0 \text{ km s}^{-1}$, and $\delta M = 0.45 \text{ km s}^{-1}$. The values of the uncertainty reported by the fitting program for the spectrum that results from combining the 5 “on-off” pairs

(shown in Figure 4.5) are $(\delta I)_{fit} = 0.008 \text{ R}$, $(\delta W)_{fit} = 3.6 \text{ km s}^{-1}$, and $(\delta M)_{fit} = 1.1 \text{ km s}^{-1}$. I infer from this exercise that the results reported by the fitting program are reasonable estimates of the uncertainties in the parameters, and therefore, that the systematic effects described above are not significantly affecting the data. I have included these values as the uncertainties in the parameters listed in Table 4.2. It must be emphasized that these parameters are those of the “on-off” spectrum and that there are also significant potential biases caused by the fact that the “off” spectrum contains some unknown quantity of $\text{H}\alpha$ emission.

Figure 4.6 shows the result of subtracting the $\text{H}\alpha$ spectrum of direction 12a from the $\text{H}\alpha$ spectrum of direction 6a. Once again an emission line is clearly seen at the same velocity as the 21 cm component. The base line is less well behaved in this case. The intensity of this line is 0.18 R , over twice the brightness found in the 1a direction. This is interesting since the H I column density at least 6 times (and possibly 25 times) greater in the 1a direction than in the 6a direction. Figure 4.2 shows that the 1a direction is located near the peak of the 21 cm enhancement while the 6a direction is at the edge of the H I cloud. The implications of this observation are discussed in more depth in Section 4.4. A more direct comparison of the $\text{H}\alpha$ intensity from the two directions could be made by subtracting the same background direction. This is impossible in this instance because one can not subtract an “off” direction spectrum taken on a different night because the sky spectrum changes and the “on-off” technique breaks down. However, There is no reason to expect large differences in the high velocity $\text{H}\alpha$ between the 4a and the 12a direction since neither has detectable 21 cm emission at the velocity of the M I cloud. The available evidence suggests that they are quite similar, making it highly implausible that the increase in

brightness seen in the 6a direction is due to the different “off” directions that were used.

Figure 4.7 shows the 2a–3a spectrum. The intensity represented by this emission line is 0.15 R, comparable to the 6a–12a spectrum. The baseline drops sharply on the red edge of the line due to bright intermediate velocity gas in the 3a direction (see below). The velocity of the $H\alpha$ emission line is once again fairly well lined up with the 21 cm emission line velocity. Since spectra of the 4a direction were obtained on three of the four observation nights, I will also use 4a as a common background direction when possible. Figure 4.8 shows the 2a–4a spectrum. The 2a spectrum is the average of spectra derived from two 900 second exposures taken on February 24, 1996. On this night we also observed the 4a direction with two 900 second exposures (see Table 4.1). In general, spectra taken on the same night are close enough in time so that the sky spectrum can be subtracted out effectively. The 2a–4a emission line has an intensity of 0.20 R. This is significantly more emission than the 2a–3a line, and the difference is mainly due to the difference in the fitted widths; the 2a–4a peak height is only slightly greater than the 2a–3a, and the blue side of the two lines look quite similar. This increase in width is a result of there being less $H\alpha$ emission from intermediate velocity gas in the 4a direction than in the 3a direction (see below). This comparison demonstrates that the results obtained for the emission lines in “on–off” pair spectra do depend somewhat on the “off” direction due to the fact that the “off” directions also contain $H\alpha$ emission. There does not appear to be much high velocity emission in our “off” directions, but there is definitely varying amounts of intermediate velocity gas emission. This is not unexpected given the ubiquity of intermediate velocity H I in this area of the sky seen in the 21 maps and the fact that the H I high velocity

clouds appear as relatively isolated enhancements.

Figure 4.9 shows the 5a–9a spectrum. This spectrum is fairly noisy, and the baseline isn’t very well behaved, but nevertheless, there appears to be an emission line at the same velocity as the 21 cm emission. The result of subtracting a 4a spectrum taken on the same night is shown in Figure 4.10. The emission line seen in the 5a–4a spectrum is somewhat fainter than the one in the 5a–9a spectrum. This is likely due to there being a small amount of high velocity $H\alpha$ emission in the 4a direction (see below). Due to the noise level and the uncertain baseline, it is even debatable if the 5a–4a spectrum shows an emission line. The fit shown in Figure 4.10 corresponds to an intensity of 0.06 R and has an uncertainty of 0.02 R. This alone would make it a 3 σ detection; however, the curve shows which data points were included in the fit and it can be seen that the fitting program’s result is not influenced by the erratic baseline seen outside the interval covered by the curve.

Figure 4.11 shows the the 8a–4a spectrum. This emission line has an intensity of 0.1 R, comparable to the 1a direction. In this case there appears to be an offset from the 21 cm component, but it is difficult to say whether this is significant. There are several factors compromising the accuracy with which we can assign velocities to the high velocity components for both the 21 cm spectra and the $H\alpha$ spectra. For one, the location of the line in the $H\alpha$ spectra is affected by the subtraction of the “off” direction. The velocity resolution of the 21 cm data is 16 km s^{-1} . Further, the components fitted in the 21 cm data are in many cases not clearly isolated lines, and this affects the accuracy with which one can assign a velocity to the component. Another factor to consider is that the 21 cm information is interpolated from the beams nearest to our beams and therefore do not sample the same sightlines.

The 7a–11a spectrum is displayed in Figure 4.12. The fact that no emission feature is seen in the spectrum provides useful information. Figure 4.2 shows that the 7a direction is just off the edge of the main H I cloud. The non-detection of $H\alpha$ in the 7a direction is therefore evidence against a large halo of ionized gas surrounding the neutral M I cloud. Figure 4.13 shows the 7a–4a spectrum. In this case, the baseline is not well determined, but again, there is no high velocity emission component; if anything, there is a dip, possibly due to the presence of a small amount of high velocity gas in the 4a direction.

A few of the $H\alpha$ observations were devoted to a search for ionized gas in the M II cloud. Figure 4.14 shows the 2b–3b spectrum, where 2b is near the peak of the M II cloud 21 cm enhancement (see Figure 4.1). The $H\alpha$ spectrum shows a clear emission component that is shifted toward less negative velocities than the M I cloud components. This velocity shift is also seen in the H I data. Figure 4.15 shows the results of subtracting the spectrum of our standard background direction, 4a, from the 2b spectrum. In this case, the emission was fitted with two Gaussian components, since the component has a non-Gaussian flat-top shape and the red side of the emission is closer to the velocity normally considered “intermediate velocity” gas. The component with a more negative velocity in this fit is centered at -78 km s^{-1} and has an intensity of 0.16 R. This velocity is not too far from the 21 cm velocity of -86 km s^{-1} found for the 2b direction. Once again, because of the difficulties in isolating the emission components it is not clear whether this discrepancy is real. Also, it is interesting to note that our results are in good agreement with the lower signal to noise Münch and Pitz (1989) detection of $H\alpha$ emission in a sightline ($l = 185.0^\circ$, $b = 65.0^\circ$) very near ours. They measured a component at $v_{lsr} = -80 \text{ km s}^{-1}$ with $I_\alpha = 0.15 \text{ R}$.

Figure 4.16 shows the 1b–4a H α spectrum. There is a clear emission component, but it is at fairly low velocity and may not even qualify as high velocity gas. It is consistent with the lower negative velocity of the 2b direction, suggesting perhaps a trend toward lower velocities for the M II cloud. There is also a hint of a weak higher negative velocity component in this spectrum near $v_{lsr} = -115 \text{ km s}^{-1}$; no 21 cm emission feature was found at this velocity.

4.3.2 Subtracting the Lockman Window Spectrum

In attempting to measure faint H α emission from the ISM, a very powerful tool would be a direction in the sky where it was known that the H α emission was zero at all velocities. Using this as the “off” direction would allow one to subtract the sky spectrum and correct for other systematic uncertainties without introducing the uncertainty caused by an unknown quantity of emission from the “off” direction. We believe that the Lockman Window direction (our 10a) forms a good approximation to this ideal. This direction has a famously low H I column density ($N_{HI} = 4.4 \pm 0.5 \times 10^{19} \text{ cm}^{-2}$; Jahoda, Lockman, & McCammon 1990). As a result, this direction is often used by X-ray astronomers and others interested in looking at sources unobscured by the affects of neutral hydrogen and the associated dust. An important question is whether the Lockman Window is similarly devoid of emission from ionized hydrogen. Because of baseline uncertainties, we cannot claim that there is zero H α emission from the Lockman Window direction; however, it appears to be fainter in H α than any other direction we have observed. We see no evidence for H α emission from the Lockman Window except for a possible very weak high velocity component at $v_{lsr} \simeq -130 \text{ km s}^{-1}$. We will eventually be able either to detect H α or put very low upper limits on

its intensity by using the shifts in velocity caused by the earth's motion to distinguish between any actual emission and baseline irregularities. At present, we treat the 10a direction as our best approximation to zero galactic emission, and attempt thereby to measure the absolute quantity of $H\alpha$ from both the high and the intermediate velocity gas from the other directions, without the biases caused by emission from an "off" direction. At this point, this technique should be classified as experimental.

Unfortunately, we only have one observation of the 10a direction, which took place on March 23, 1996. Since we know the sky spectrum changes from night to night, we cannot simply subtract the 10a spectrum from other nights. Table 4.1 shows that the March 23 observations include 1a, 4a, 5a, 9a, 7a, 11a, and 8a. We developed a technique whereby we can transform a spectrum from one night to what it would have looked like on another night in cases where the same galactic direction was observed on both nights. The difference between the 1a observation on 3/23/96 and the one on 2/24/96, since the galactic emission is the same in both cases, is just the difference in the sky spectrum between these two nights. Adding this difference to a spectrum of 4a taken on 2/24/96 will produce what the spectrum of 4a would have looked like if taken on 3/23/96. Since both 1a and 4a were taken on both dates, I was able to verify that this technique works very well. This technique made it possible to subtract the 10a spectrum from 2a, 3a, 1b, 2b, and 3b from the 2/24/96 observations. A few examples of the resulting Lockman Window subtracted spectra are discussed below.

Figure 4.17 shows the 1a-10a spectrum. In this case, we can clearly see a separate emission component from the high velocity gas in the M I cloud as well as a general trend toward more emission at lower negative velocity. There is perhaps also

a separate intermediate velocity component, but this is less certain because geocoronal subtraction problems can dominate at velocities more positive than -50 km s^{-1} . Figure 4.18 shows the 4a–10a spectrum. The intermediate velocity emission from this direction is very similar to that seen in 1a and this explains why the baseline is so well behaved in the 1a–4a spectrum. There is a hint of a high velocity component ($\lesssim 0.02 \text{ R}$) in the 4a–10a spectrum at the same velocity as the 1a component. Figure 4.19 and Figure 4.20 show the 2a–10a spectrum and the 3a–10a spectrum, respectively. Strong $\text{H}\alpha$ emission is seen at high velocity from the 2a direction. The 3a direction shows stronger intermediate velocity emission, which explains the sharp drop on the red side of the 2a–3a emission feature. Finally, Figure 4.21 shows the 8a–10a direction. This is a good example of a case where the high velocity emission blends smoothly into the intermediate velocity gas. In some cases, it is only the similarity of the intermediate velocity gas in the “on” and the “off” directions that allows the extraction of separated high velocity emission lines.

4.3.3 The [S II] $\lambda 6716$ “on–off” Spectra

Figure 4.22 shows the spectrum of the 1a–4a direction in the [S II] $\lambda 6716$ line. The spectrum shows a fairly solid detection of this line at a velocity very close to the velocity at which $\text{H}\alpha$ emission was detected in this direction, which is indicated by an arrow in the figure. The intensity of this line is $0.05 \pm 0.01 \text{ R}$, and the fitted width is $23 \pm 8 \text{ km s}^{-1}$ (the uncertainties quoted here are the formal errors calculated in the fitting process). The mean velocity determined from the fit is $v_{lsr} = -110 \pm 3 \text{ km s}^{-1}$. The spectrum also shows another little peak near -60 km s^{-1} . Figure 4.23 shows the [S II] $\lambda 6716$ spectrum of the 6a–12a direction. The spectrum does not show any clear

detections; however, there is a marginal enhancement at -96 km s^{-1} . The Gaussian fit to this enhancement corresponds to an intensity of $0.02 \pm 0.01 \text{ R}$ and has a width of 18 km s^{-1} . I will treat the fitted intensity as an upper limit to the emission from this direction. The $[\text{S II}] / \text{H}\alpha$ ratios for the two directions are then:

$$\left(\frac{I_{\text{S II}}}{I_{\text{H}\alpha}} \right)_{1a4a} \simeq 0.64 \pm 0.14, \quad (4.1)$$

and

$$\left(\frac{I_{\text{S II}}}{I_{\text{H}\alpha}} \right)_{6a12a} \simeq 0.11 \pm 0.06, \quad (4.2)$$

a difference of a factor of six! The cause of this drastic gradient in the $[\text{S II}] / \text{H}\alpha$ ratio is not known, but it clearly indicates that something is very different between the ionization/excitation conditions in these two directions.

4.4 Discussion and Conclusions

The above results allow some general conclusions to be drawn about the pattern of $\text{H}\alpha$ intensities associated with these high velocity clouds. First, in every direction where there is significant high velocity H I gas seen in 21 cm emission, we have found associated ionized hydrogen emitting the $\text{H}\alpha$ line. Second, there is no clear evidence of $\text{H}\alpha$ in any of the directions without appreciable H I emission. Nevertheless, the intensities of the 21 cm emission and the $\text{H}\alpha$ emission are not well correlated. Figure 4.24 shows the $\text{H}\alpha$ intensity in Rayleighs plotted versus the 21 cm intensity in brightness temperature units. The $\text{H}\alpha$ intensities are from the fits to the “on-off” spectra, where 4a was used as a background direction, except in the case of 6a, where 12a is used as the “off”. The intensity for the 2b direction is from the high velocity component in the two Gaussian fit. No correlation is apparent in the plot. If anything, the two

intensities appear to be anticorrelated; however, the impression of an anticorrelation derives from the one or two data points with the highest 21 cm intensity, and the data sample is too small to derive any clear conclusions about an anticorrelation. There is some evidence that the ionized gas producing the $H\alpha$ emission envelopes the 21 cm emitting neutral gas (the brightness of the 6a direction in $H\alpha$) but the $H\alpha$ “halo”, if present, is not large (for example, the nondetection of $H\alpha$ in the 7a direction). A complete map of the high velocity $H\alpha$ emission in this region, which we hope to produce after the WHAM survey is finished, would be very illuminating in this regard.

The pattern of $H\alpha$ intensities does not seem to correlate well with the soft X-ray brightness derived by Herbstmeier et al. (1995). They calculate that there is an enhancement in soft X-rays emitted by the M I cloud in direction 1a. We do not see enhanced $H\alpha$ in this direction. Direction 5a, with no strong X-ray enhancement on the Herbstmeier et al. map, has both comparable 21 cm emission and comparable $H\alpha$ emission (see Table 4.2). On the M II cloud, the 1b direction is at the location of another of the Herbstmeier et al. soft X-ray enhancements, and the 2b direction is located where the soft X-ray emission is thought to be faint. We, on the other hand, find the 2b direction to be significantly brighter in $H\alpha$ than the 1b direction.

The velocities of the components found in the $H\alpha$ spectra, however, do correlate with the velocities found in the 21 cm data. Figure 4.25 shows a plot of the $H\alpha$ component velocities versus the 21 component velocities. A good correlation between the two exists, indicating that the $H\alpha$ emission is clearly associated with the H I clouds seen in the 21 cm line. The scatter is consistent with our current measurement errors. The systematic offset is tantalizing, but may well also be due to measurement error.

The widths of the $H\alpha$ components are very uncertain. This is due to the fact that

in many cases the high velocity emission is blended with intermediate velocity emission, which our “on–off” method is able to subtract out, but not without introducing biases into the measured line widths for the high velocity emission. The 1a–4a direction provides the best measurement we have of a high velocity emission component, and in this direction we measure the width to be $27 \pm 4 \text{ km s}^{-1}$. This is quite similar to the $20\text{--}25 \text{ km s}^{-1}$ line widths typical of the 21 cm emission components in this region (Giovanelli et al. 1973). The bias due to the “off” subtraction can be seen by comparing the width found for the 2a–3a spectrum ($W = 38 \text{ km s}^{-1}$) to that measured for the 2a–4a spectrum ($W = 49 \text{ km s}^{-1}$). As another example, the 5a–9a emission line has a width of 46 km s^{-1} , while the 5a–4a emission line is 38 km s^{-1} wide. Therefore, the component widths are not determined well, and the width uncertainties listed in Table 4.2, which were calculated from the quality of the fits and without considering the bias due to the “off” subtraction, are underestimated. The [S II] line width in the 1a direction appears to be narrower than the $\text{H}\alpha$ line width, but is highly uncertain. The formal fit gives $23 \pm 8 \text{ km s}^{-1}$ (FWHM) for the [S II] line width compared to $27 \pm 4 \text{ km s}^{-1}$ for the $\text{H}\alpha$ line.

Since interstellar extinction can be neglected along these high latitude sightlines, the $\text{H}\alpha$ intensity towards a high-velocity cloud is directly related to the emission measure ($\text{EM} \equiv \int n_e^2 dl$) through the cloud and gives information on the column density of ionized hydrogen and the electron density associated with the cloud. As an example, consider the 21 cm enhancement centered near direction 1a. It has a peak H I column density of $2 \times 10^{20} \text{ cm}^{-2}$ and a diameter of about 1.5° , which corresponds to 50 pc at a distance of 2 kpc. Hence the density of neutral hydrogen $n_{\text{H I}} \simeq 1.3 \text{ cm}^{-3}$, and the mass of the cloud is about $2000 M_\odot$. For direction 1a, we measured I_α to be

approximately 0.08 R, implying an emission measure $EM = 0.18 \text{ cm}^{-6} \text{ pc}$, assuming a temperature for the H II of 8000 K. This leads to an electron density $n_e \simeq 0.06 f^{-1/2} \text{ cm}^{-3}$ and a column density of ionized hydrogen $N_{\text{H II}} \simeq 1 \times 10^{19} f^{1/2} \text{ cm}^{-2}$, where f is the filling fraction of the H II within the cloud. The H I column density in the 1a direction is $N_{\text{HI}} \simeq 1.25 \times 10^{20} \text{ cm}^{-2}$. Since $f \leq 1$ and there is little evidence that the H II associated with the cloud extends significantly beyond the H I portion of the cloud (see above), this would suggest that at this location the HVC is primarily neutral with $N_{\text{HII}} / N_{\text{HI}} \lesssim 0.08$. Interestingly, most of the H II may be located near the outer surface of the cloud. For example, if we consider direction 6a which appears to intersect the edge of the cloud (see Figure 4.1), the H I column density is less than $2 \times 10^{19} \text{ cm}^{-2}$, whereas the $\text{H}\alpha$ is twice brighter, indicating that there may be as much or more H II as H I along the edge of the H I enhancement.

If the $\text{H}\alpha$ emission arises from the photoionization of the cloud, then the intensity of the emission is directly related to the incident Lyman continuum flux F_{LC} , such that

$$F_{LC} = 1.7 \times 10^5 \left(\frac{I_\alpha}{0.1R} \right) \text{ photons cm}^{-2}\text{s}^{-1}.$$

If, on the other hand, the ionization is due to a shock arising from the collision of the high-velocity gas with an ambient medium in the halo, then we can gain information on the density and temperature of this ambient medium. A cloud velocity of 100 km s^{-1} is supersonic for ambient temperatures $T \leq 10^6 \text{ K}$. If the ambient gas is sufficiently cool for strong shocks to occur, then we can use shock models by Raymond (1979), which apply to shocks with $50 \leq V_s \leq 140 \text{ km s}^{-1}$, to relate the face-on $\text{H}\alpha$

surface brightness $(I_\alpha)_\perp$ to the halo gas parameters,

$$(I_\alpha)_\perp \simeq 6.5 n_o \left(\frac{V_s}{100} \right)^{1.6} R,$$

where n_o is the density of the preshocked gas. Hence, if $V_s = 100 \text{ km s}^{-1}$, and $(I_\alpha)_\perp = 0.1 R$, then $n_o \leq 0.015 \text{ cm}^{-3}$. This is an upper limit on n_o because a nonperpendicular sightline will increase the observed I_α for a given n_o and V_s .

Perhaps the important question of the source of ionization can be explored in the future by measuring various emission line ratios. Models by Raymond (1979) predict optical line ratios for shocks under various conditions. With an ambient density $n_o = 1 \text{ cm}^{-3}$ and a shock velocity $V_s = 100 \text{ km s}^{-1}$, some of the predicted line ratios are $[\text{S II}] \lambda 6716 / \text{H}\alpha = 0.18$, $[\text{N II}] \lambda 6584 + \lambda 6548 / \text{H}\alpha = 0.51$, and $[\text{O III}] \lambda 5007 + \lambda 4959 / \text{H}\alpha = 1.9$. If this model, when extended to lower densities, retained the prediction of very bright $[\text{O III}]$ emission, then the $\lambda 5007$ line could be an important discriminator between shock ionization and photoionization. A more careful map of I_α around the edge of the clouds could also be used to distinguish between the two ionization mechanisms.

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Name	l (°)	b (°)	Exp. Time (seconds)	Obs. Date †
1a	168.30	65.20	4500	1,2,3
2a	163.30	66.70	1800	2
5a	165.80	65.95	900	3
6a	170.85	64.72	900	4
7a	168.30	67.82	900	3
8a	163.30	64.70	900	3
1b	183.40	65.75	900	2
2b	186.00	65.00	900	2
4a	168.00	70.93	4500	1,2,3
3a	168.30	69.00	1800	2
9a	167.00	60.00	900	3
12a	169.00	74.00	900	4
11a	140.00	80.00	900	3
10a	150.48	52.96	900	3
3b	186.0	70.00	900	2

Table 4.1: HVC H α Observation Data.

† Observations were carried out on 4 nights, 1: 2/18/96, 2: 2/24/96, 3: 3/23/96,
4: 4/10/96.

Name	$H\alpha$						21 cm $\ddagger$	
	$I_{H\alpha}$ (R)	$\delta I_{H\alpha}$ (R)	$V_{H\alpha}$ (km s $^{-1}$)	$\delta V_{H\alpha}$ (km s $^{-1}$)	$W_{H\alpha}$ (km s $^{-1}$)	$\delta W_{H\alpha}$ (km s $^{-1}$)	T_B (K)	$V_{H\ I}$ (km s $^{-1}$)
M I								
1a-4a	0.08	0.01	-106	1	27	4	2.59	-112
2a-3a	0.15	0.02	-98	1	38	3	1.19	-99
2a-4a	0.20	0.02	-95	1	49	4	1.19	-99
5a-9a	0.09	0.02	-109	2	46	8	1.49	-116
5a-4a	0.06	0.02	-103	3	38	10	1.49	-116
6a-12a	0.18	0.02	-105	1	35	4	0.11	-108
7a-11a	—	—	—	—	—	—	—	—
8a-4a	0.11	0.02	-92	2	50	7	0.71	-107
M II								
1b-4a	0.12	0.01	-61	2	44	5	—	—
2b-3b	0.20	0.02	-72	1	44	4	0.25	-86
2b-4a	0.16	0.03	-78	2	27	5	0.25	-86

Table 4.2: HVC $H\alpha$ Results. $\ddagger$ H I data from Bart Wakker.

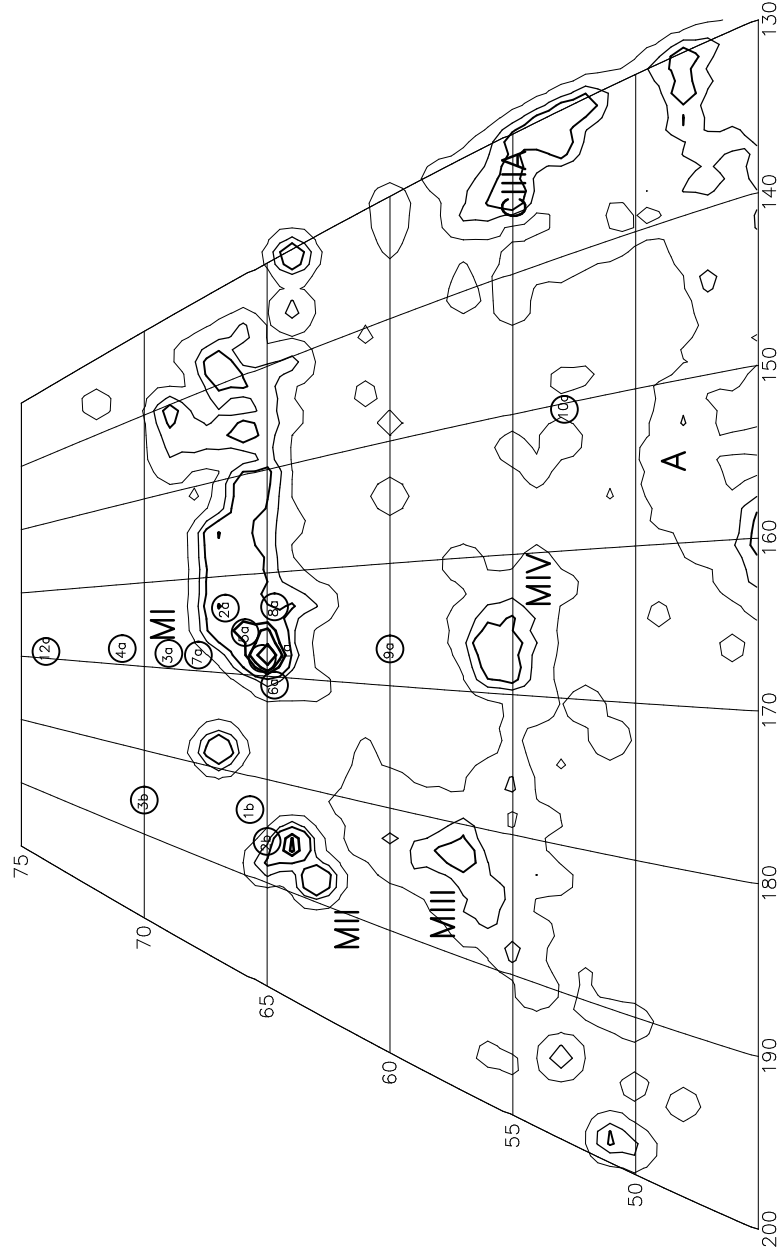


Fig. 4.1.— $H\alpha$ look directions on a map showing 21 cm contours. The map and the H I data were provided by Bart Wakker.

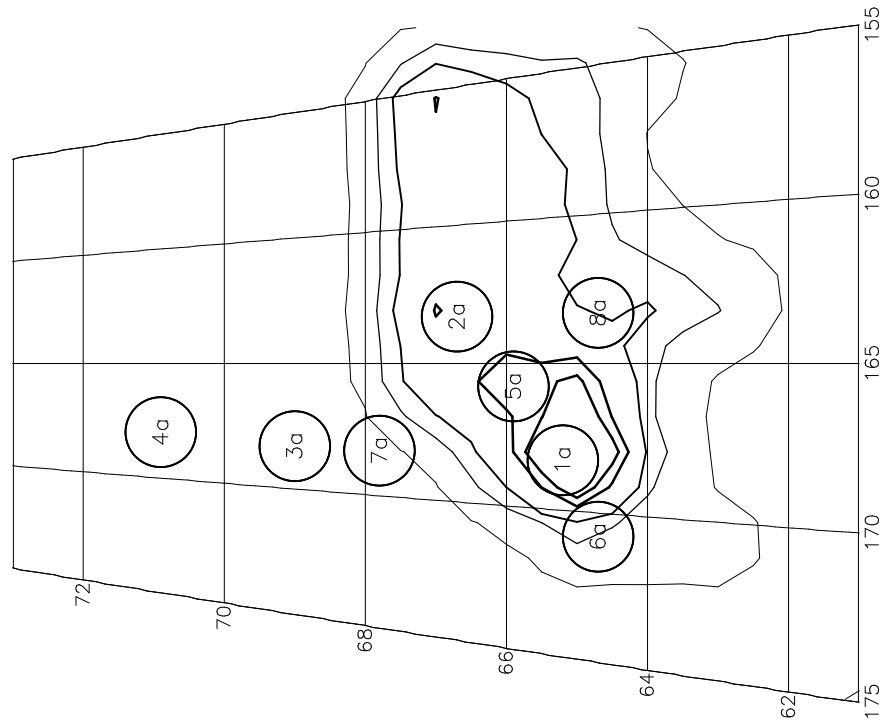


Fig. 4.2.— $H\alpha$ look directions on the M I cloud. The 21 cm map and data were provided by Bart Wakker.

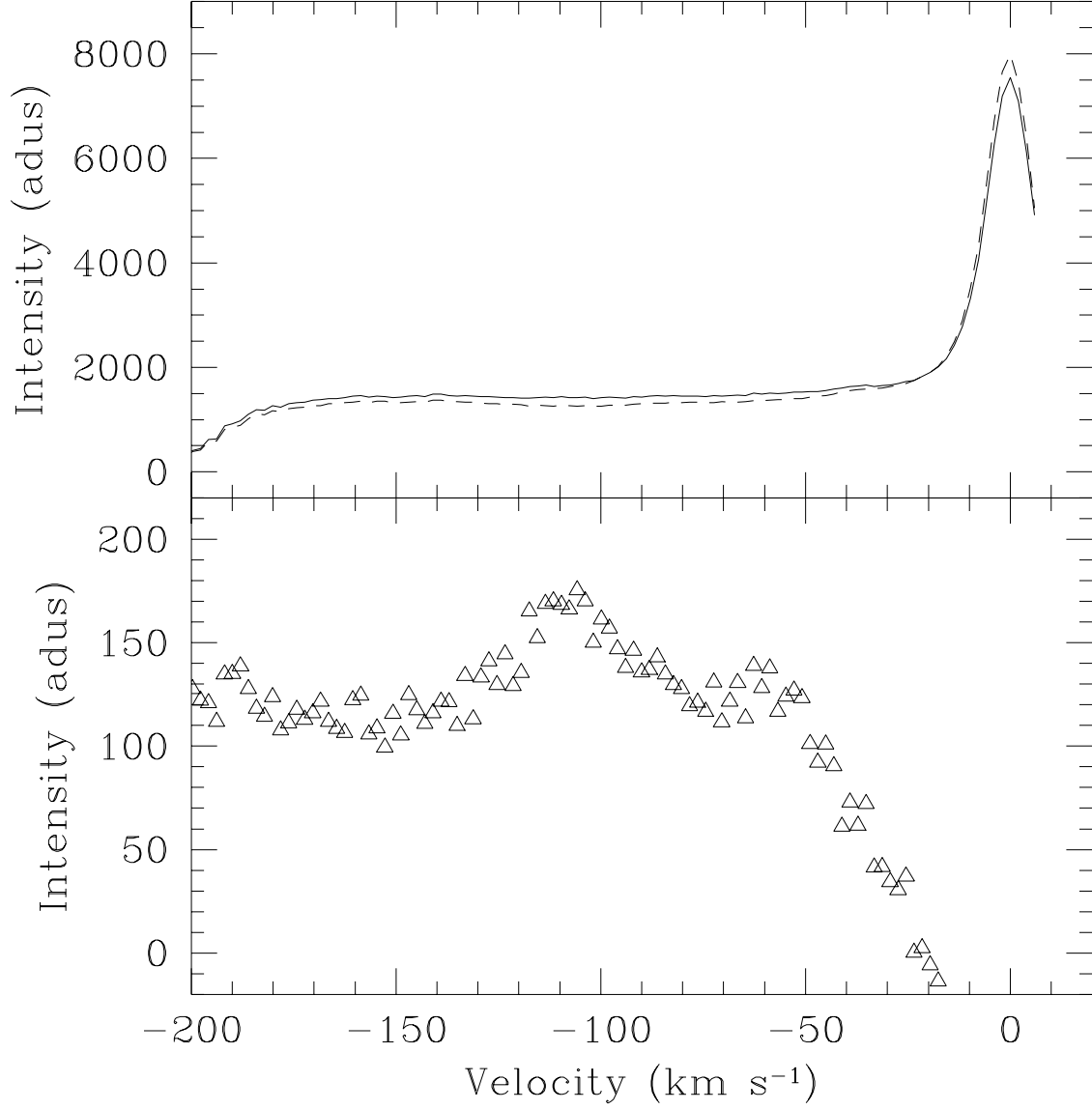


Fig. 4.3.— $\text{H}\alpha$ from the M I cloud: the upper panel shows the spectra derived from single 900 second exposures “on” (the solid line is direction 1a) and “off” (dashed line is direction 4a) the M I cloud. The lower panel shows the difference (“on - off”) spectrum after flat fielding.

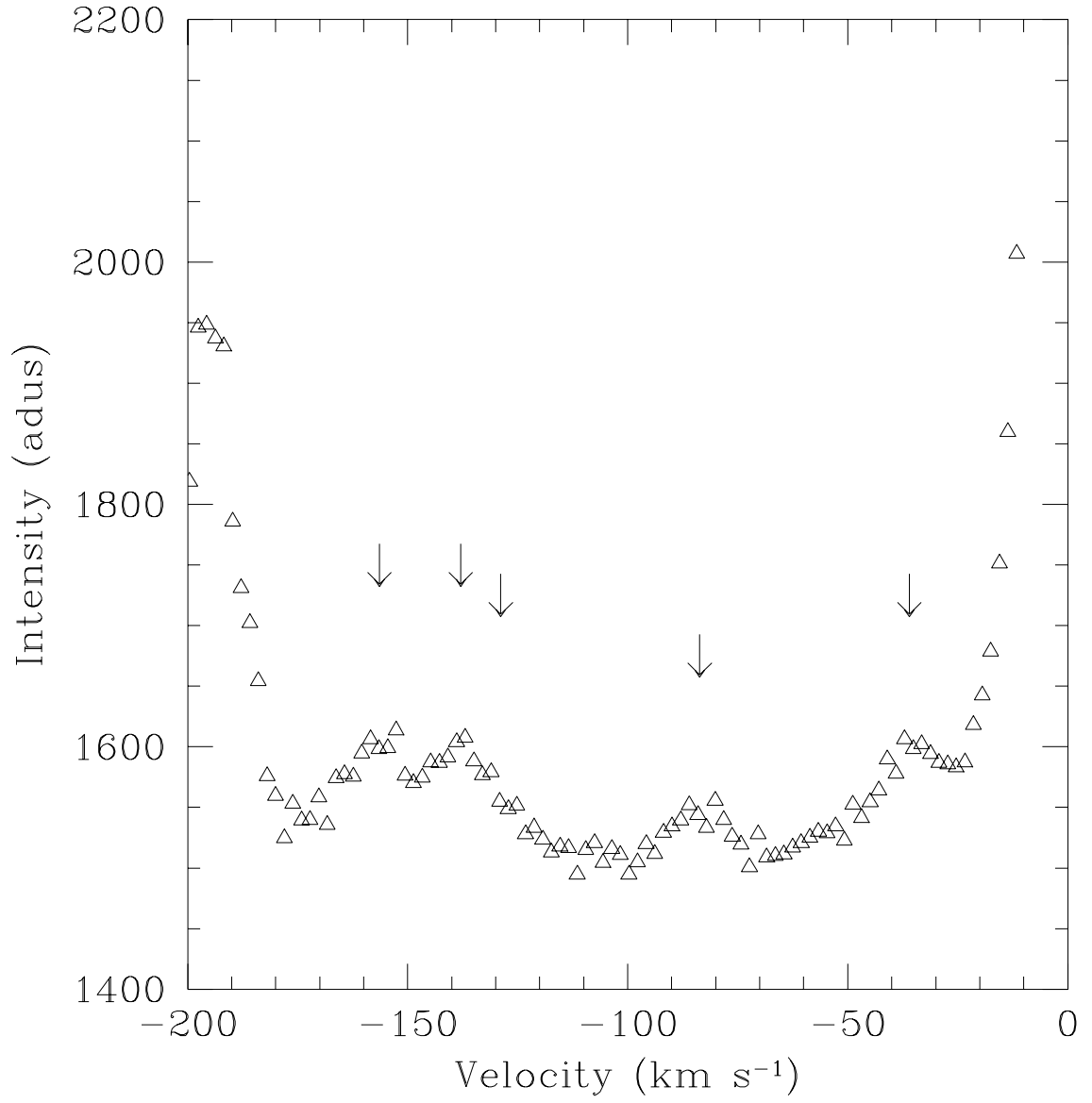


Fig. 4.4.— Night sky spectral features that are present with variable strength in all the HVC spectra.

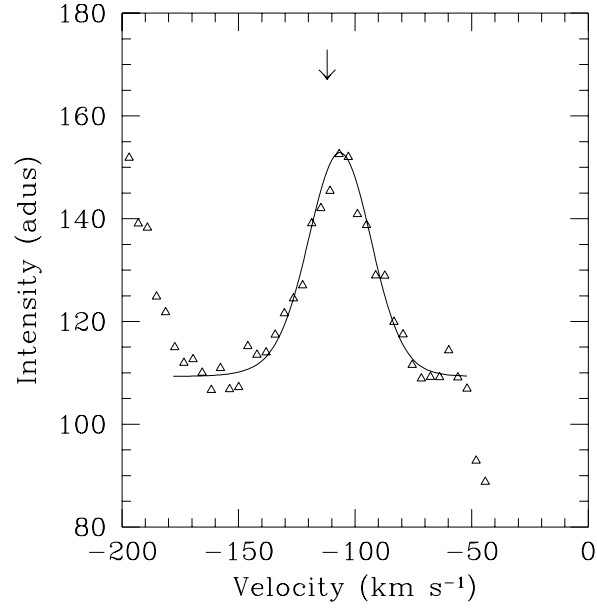


Fig. 4.5.— $H\alpha$ from the M I cloud: 1a - 4a (normalized to 900s exposure time). The arrow shows the velocity of the 21 cm HVC detection.

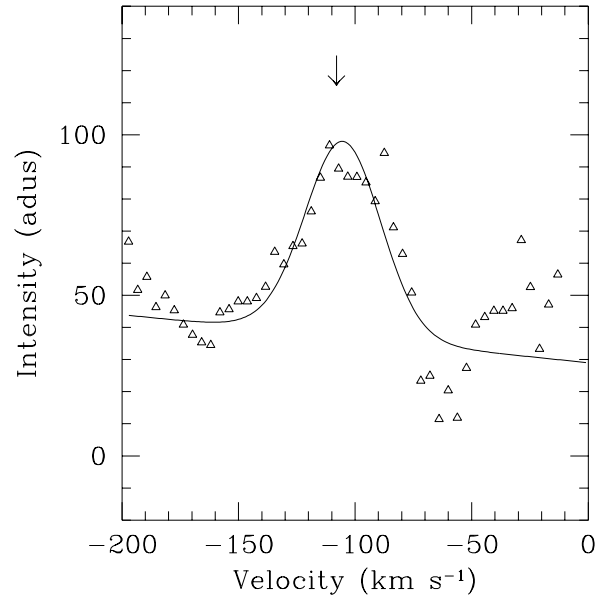


Fig. 4.6.— $H\alpha$ from the M I cloud: 6a - 12a (900s exposure). The arrow shows the velocity of the 21 cm HVC detection.

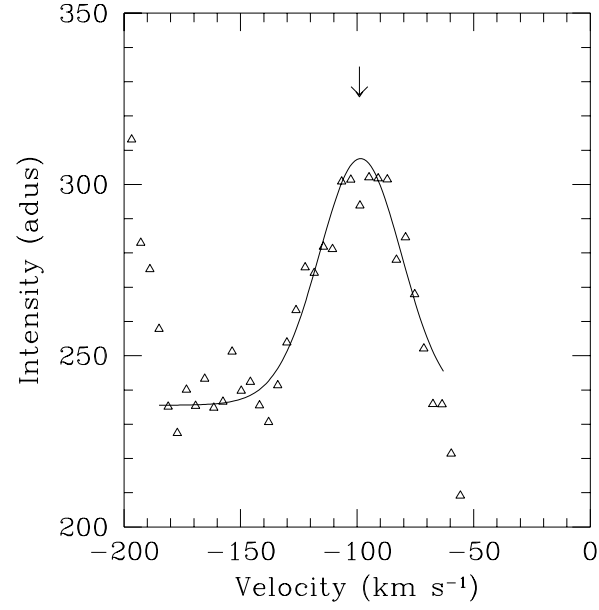


Fig. 4.7.— $H\alpha$ from the M I cloud: 2a - 3a (normalized to 900s exposure). The arrow shows the velocity of the 21 cm HVC detection.

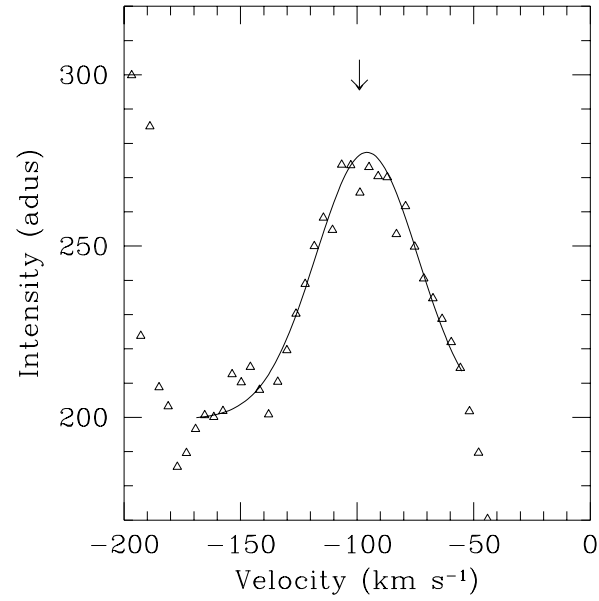


Fig. 4.8.— $H\alpha$ from the M I cloud: 2a - 4a (normalized to 900s exposure). The arrow shows the velocity of the 21 cm HVC detection.

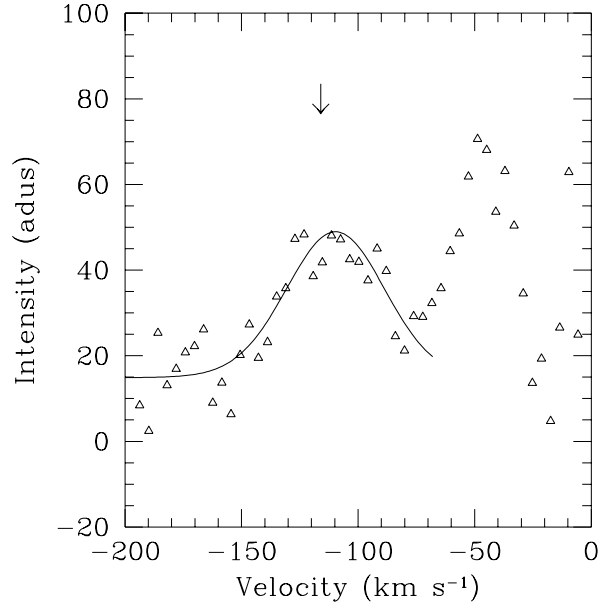


Fig. 4.9.— $H\alpha$ from the M I cloud: 5a - 9a (900s exposure). The arrow shows the velocity of the 21 cm HVC detection.

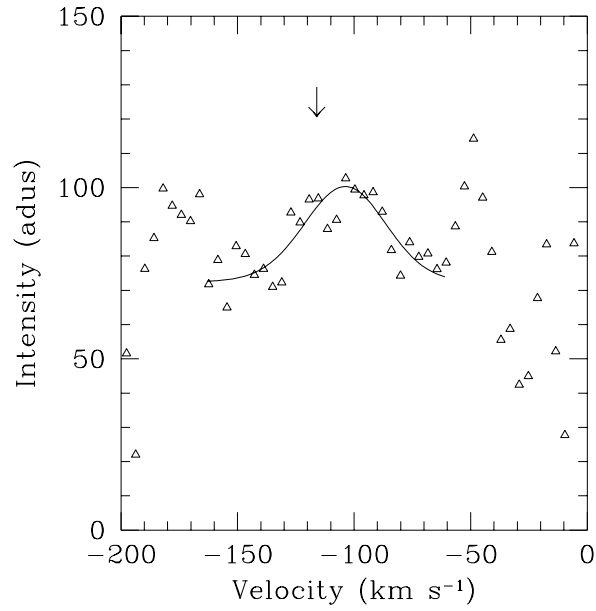


Fig. 4.10.— $H\alpha$ from the M I cloud: 5a - 4a (900s exposure). The arrow shows the velocity of the 21 cm HVC detection.

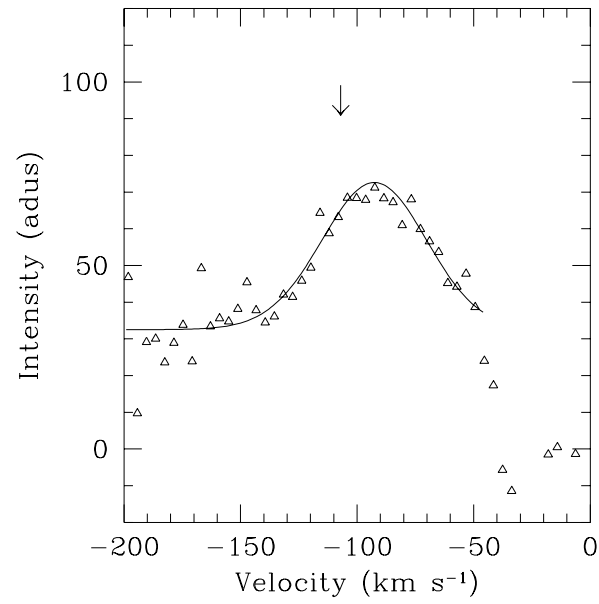


Fig. 4.11.— $H\alpha$ from the M I cloud: 8a - 4a (900s exposure). The arrow show the velocity of the 21 cm HVC detection.

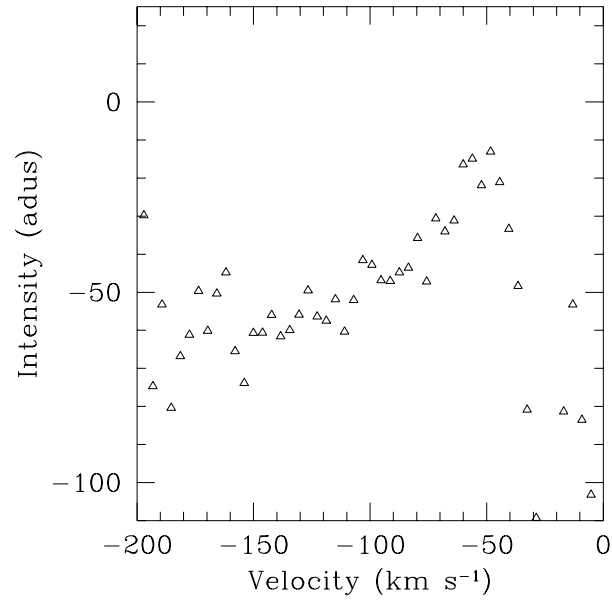


Fig. 4.12.— H α from the M I cloud: 7a - 11a (900s exposure)

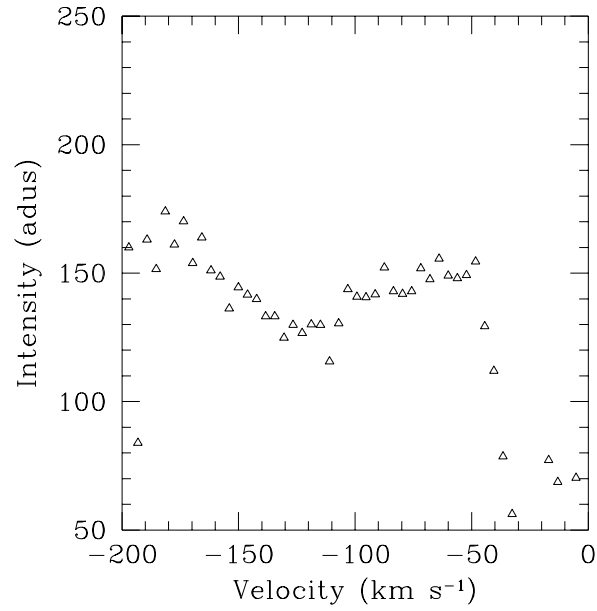


Fig. 4.13.— H α from the M I cloud: 7a - 4a (900s exposure).

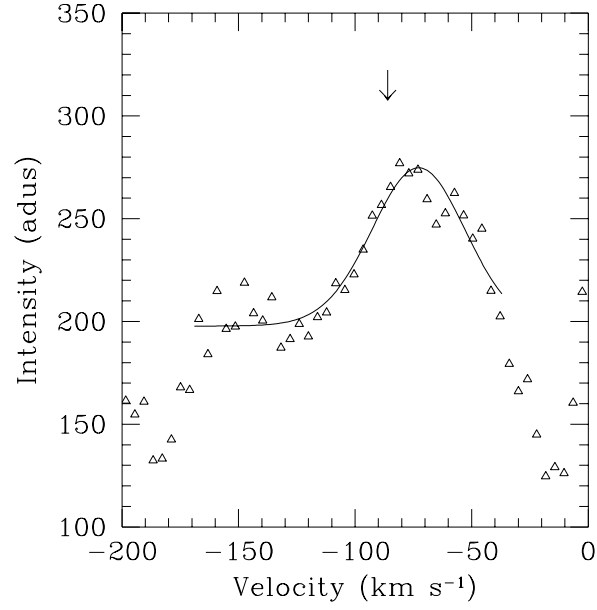


Fig. 4.14.— $H\alpha$ from the M I cloud: 2b - 3b (900s exposure). The arrow shows the velocity of the 21 cm HVC detection.

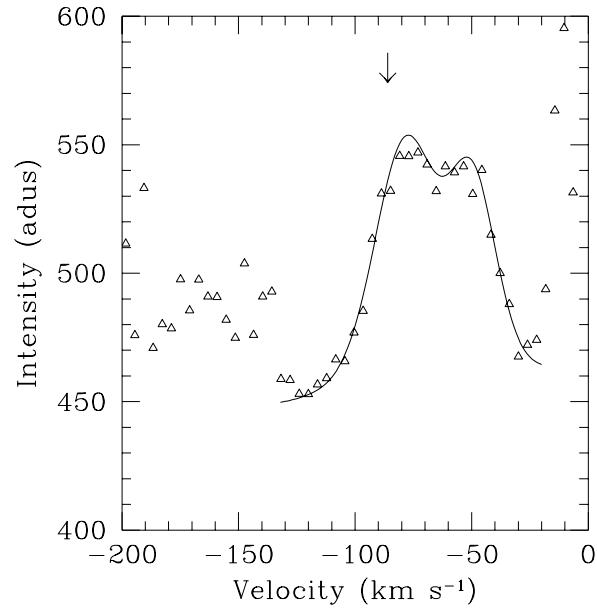


Fig. 4.15.— $H\alpha$ from the M I cloud: 2b - 4a (900s exposure). The arrow shows the velocity of the 21 cm HVC detection.

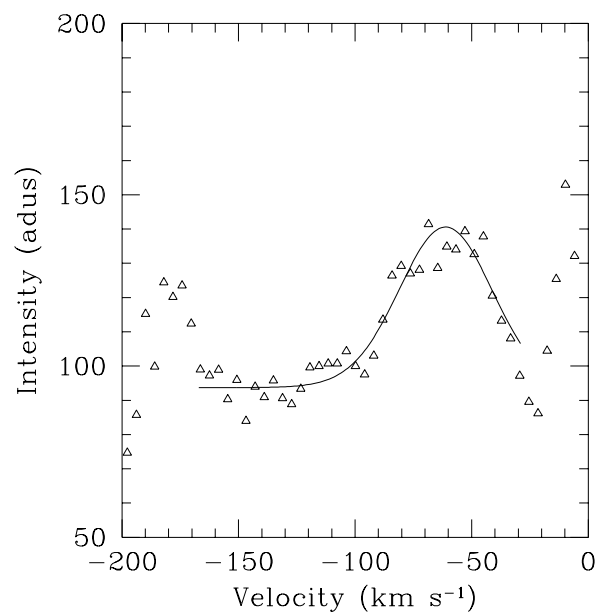


Fig. 4.16.— H α from the M I cloud: 1b - 4a (900s exposure).

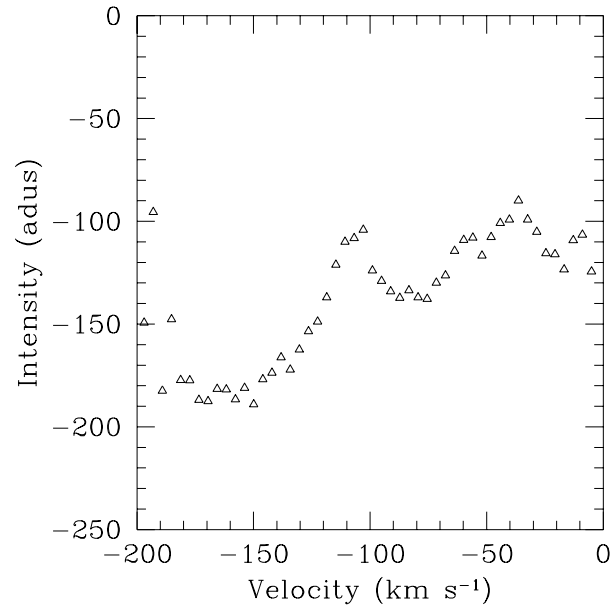


Fig. 4.17.— H α from the M I cloud: 1a - 10a (normalized to 900s exposure).

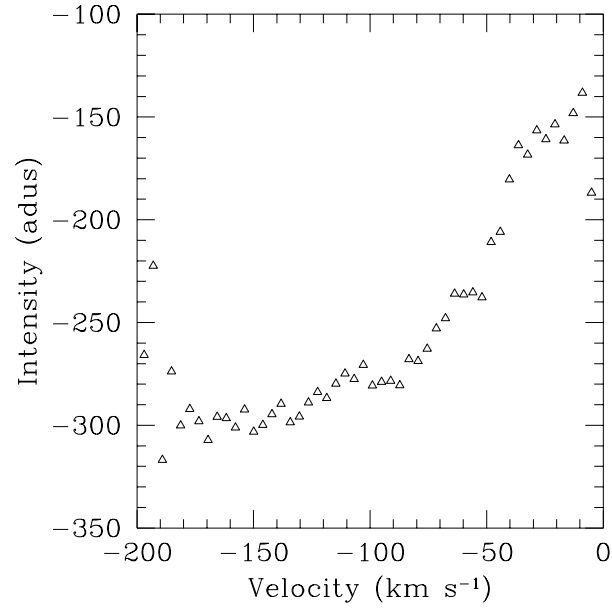


Fig. 4.18.— H α from the M I cloud: 4a - 10a (normalized to 900s exposure).

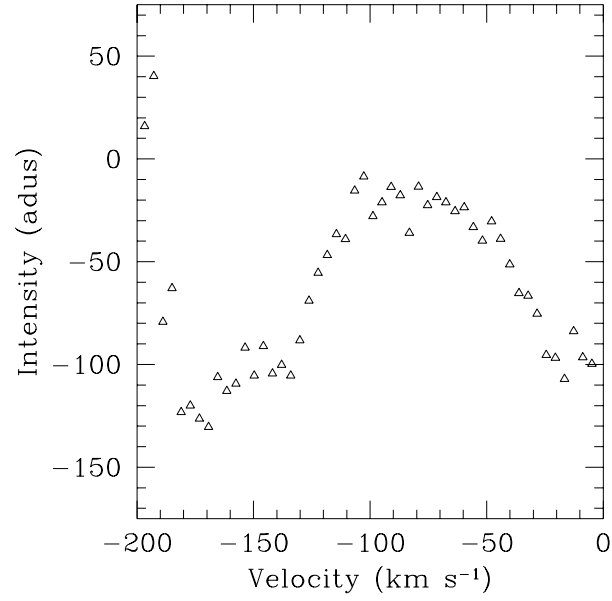


Fig. 4.19.— $\text{H}\alpha$ from the M I cloud: 2a - 10a (normalized to 900s exposure).

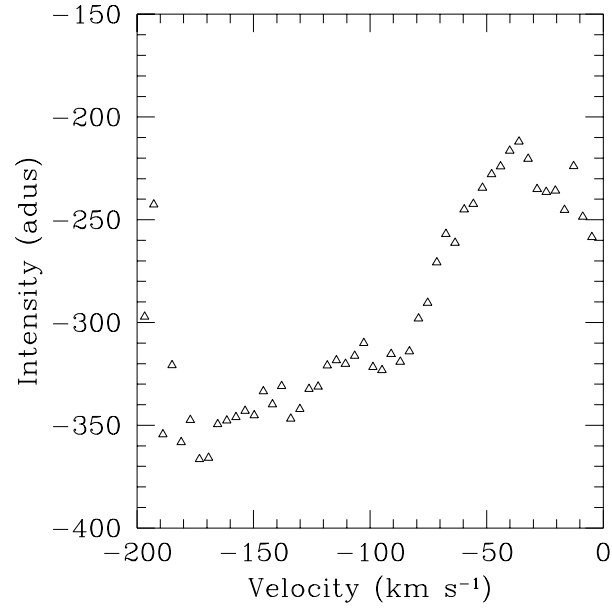


Fig. 4.20.— $\text{H}\alpha$ from the M I cloud: 3a - 10a (normalized to 900s exposure).

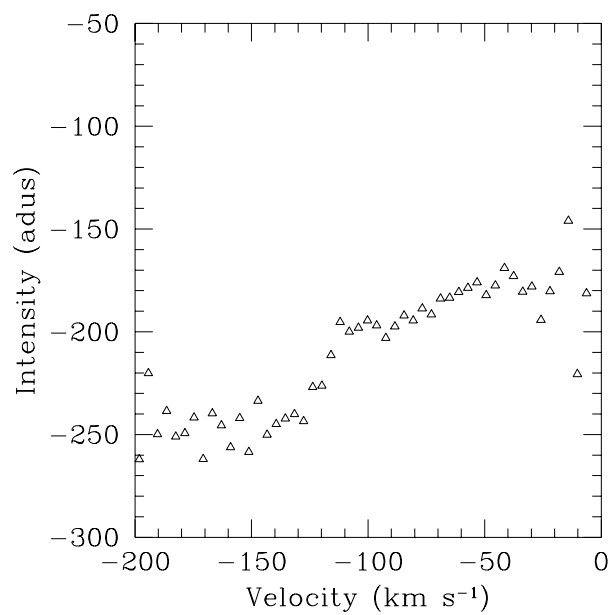


Fig. 4.21.— H α from the M I cloud: 8a - 10a (900s exposure).

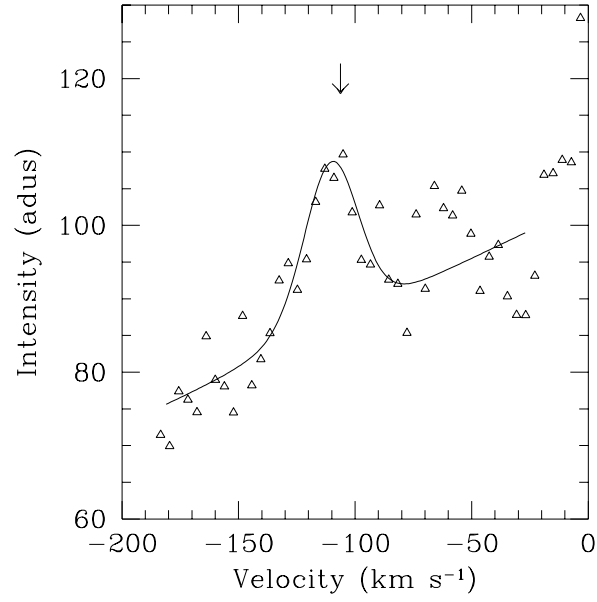


Fig. 4.22.— [S II] from the M I cloud: 1a - 4a (normalized to 900s exposure). The arrow shows the velocity of the $H\alpha$ emission line.

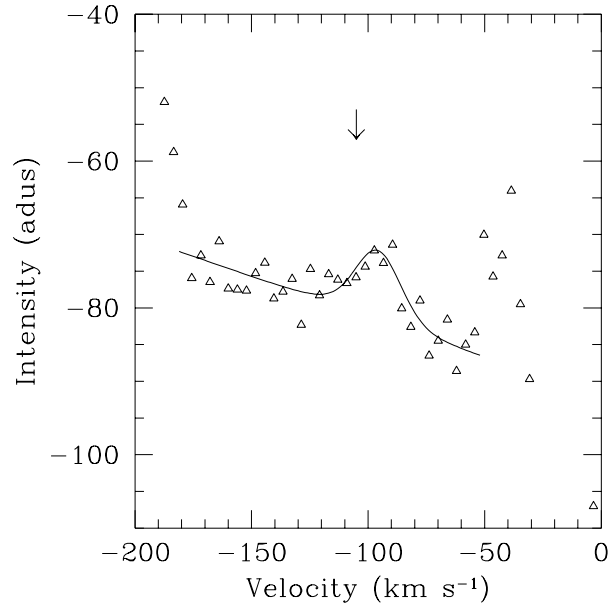


Fig. 4.23.— [S II] from the M I cloud: 6a - 12a (normalized to 900s exposure). the arrow shows the velocity of the $H\alpha$ emission line.

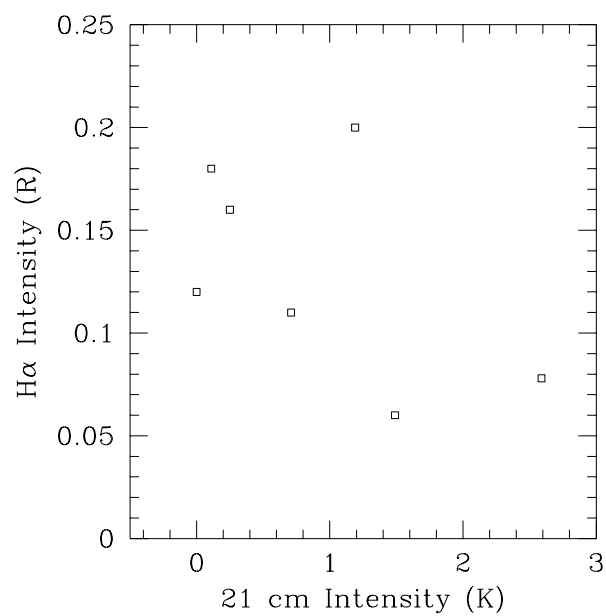


Fig. 4.24.— H α intensity versus 21 cm intensity.

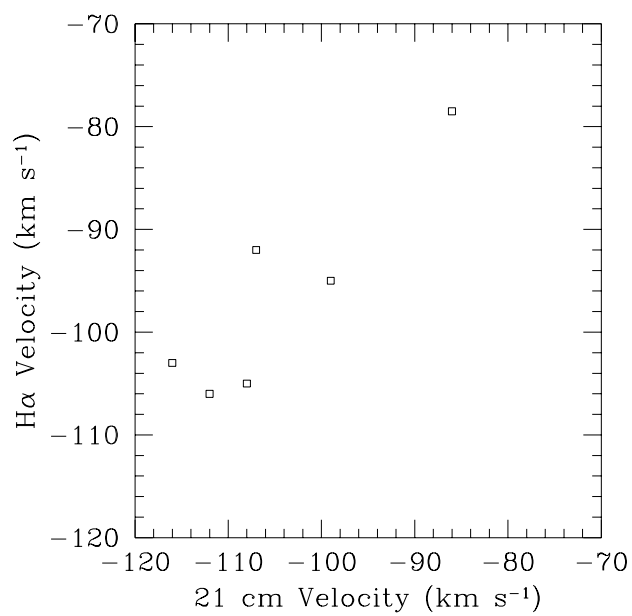


Fig. 4.25.— H α velocity versus 21 cm velocity.

Chapter 5

Detection of the He I $\lambda 5876$

Recombination Line from the WIM with the WHAM Spectrometer

5.1 Introduction

The warm ionized medium (WIM) has been shown to be a major component of the interstellar medium. Maintaining the ionization of the WIM requires a total power equal to about 15% of the ionizing photons emitted by O stars or all of the kinetic energy injected by supernova; yet the source of ionization is still not known. Recent measurements by Reynolds & Tufte (1995), and by Heiles et. al. (1996) of the fractional ionization of helium with respect to hydrogen in the diffuse interstellar medium in the Galaxy have challenged models which consider O stars the primary source of ionization for the WIM. The low fractional ionization of helium with respect to hydrogen imply that the spectrum of the radiation ionizing the WIM is softer than the spectrum produced by O stars. Rand (1997) has found a similar problem for the

diffuse ionized gas (DIG) in NGC 891, a nearby edge-on spiral galaxy. We have used the increased sensitivity of the WHAM spectrometer to obtain the first detections of the He I $\lambda 5876$ line from the WIM. The strength of this recombination line with respect to the $H\alpha$ intensity measures the fractional ionization of helium with respect to hydrogen.

5.2 Observations and Results

The $\lambda 5876$ line is the brightest recombination line of He I in the visible region and results from the $^3D \rightarrow ^3P$ ($n = 3-2$) transition. The relative emissivity of this line to $H\alpha$ in low density regions can be expressed as

$$\frac{E_{5876}}{E_{H\alpha}} \simeq 0.048 \frac{\chi_{He}}{\chi_H} \left(\frac{He/H}{0.1} \right) \left(\frac{T}{8000} \right)^{-0.14}, \quad (5.1)$$

where χ_{He} and χ_H denote the fraction of He and H, respectively, that is singly ionized, He/H is the abundance of helium with respect to hydrogen by number, T is the electron temperature, and E is the emissivity in $\text{ergs cm}^{-3} \text{ s}^{-1}$. Equation 5.1 is an analytic approximation to the results in Brockelhurst (1972) and Martin (1988) valid for $2500 \text{ K} \leq T \leq 20000 \text{ K}$. Note that this relationship was stated incorrectly in Reynolds & Tufte (1995). The constant was listed as 0.037 instead of 0.048, but the results of the paper were unaffected by the error since the constant divided out in the method used. To estimate the expected brightness of the He I $\lambda 5876$ line, consider that the brightest $H\alpha$ intensity measured from the warm ionized medium is $\simeq 10 \text{ R}$; therefore, if helium were ionized to the same degree as hydrogen in the WIM, then the brightest directions would have $I_{5876} \simeq 0.5 \text{ R}$.

We measured the He I $\lambda 5876$ intensity and the $H\alpha$ intensity in four directions

in the Galactic plane. The directions are $l = 194^\circ, 186^\circ, 114^\circ$, and 96° , with $b = 0^\circ$ in all cases. These directions were chosen based on the fact that they are bright in $H\alpha$, they do not intersect any known H II regions, and they have high $[S\ II] / H\alpha$ ratios. This last criterion is used because high $[S\ II] / H\alpha$ ratios are a signature of WIM emission and indicate that the emission is from a truly diffuse source, and not unresolved classical H II regions (Reynolds 1985). In order to reduce the confusing effects of weak atmospheric emission lines and faint “ghosts” of brighter atmospheric lines that leak through other orders of the Fabry-Perot spectrometer, we use the same “on-off” technique that was described in Chapter 4. The “off” directions were chosen at higher galactic latitudes where the $H\alpha$ intensity is much fainter.

We also measured the He I $\lambda 5876$ intensity and the $H\alpha$ intensity toward several “bright” O star H II regions and express our WIM results with respect the $I_{5876} / I_{H\alpha}$ measured for the H II regions. This eliminates problems associated with differences in the sensitivity of the WHAM system to different wavelengths and due to interstellar extinction (the H II regions appear to experience approximately the same extinction as the “background” directions [Reynolds & Tufte 1995]). We intend to make measurements of the sensitivity of the WHAM system as a function of wavelength using standard stars (see Section 3.7), but we have not done this yet. Table 5.1 shows the H II regions that were measured. The results for the $I_{5876} / I_{H\alpha}$ ratio in the H II regions indicate that at present we can only determine this ratio to an accuracy of about 35%. Separate observations of the $I_{5876} / I_{H\alpha}$ ratio for NAN yield results which differ by up to 17%, significantly more than our anticipated measurement uncertainties. For the Rosette Nebula, observations carried out on two different nights led to $I_{5876} / I_{H\alpha}$ ratios differing by $\simeq 35\%$. The reason for this lack of reproducibility is unknown at

present. Until further studies can be done to discover the source of this variation, we will be forced to accept this variation as part of our uncertainty. Understanding this problem takes a high priority since the additional uncertainty will limit the strength of the conclusions that we can draw from our detections of He I $\lambda 5876$ from the WIM.

Let χ_{He} and χ_H denote the fractional ionization of helium and hydrogen, respectively (i.e., $\chi_{He} = (\text{He})^+ / (\text{He})_{tot}$, and $\chi_H = \text{H}^+ / \text{H}_{tot}$.) If we divide our measured $I_{5876} / I_{H\alpha}$ from the WIM by the same ratio measured in the bright H II regions, it can be seen from Equation 5.1 that the result is equal to $(\chi_{He}/\chi_H)_{WIM} / (\chi_{He}/\chi_H)_{HII}$. Hence, assuming that both helium and hydrogen are fully ionized in the H II regions, that is, $(\chi_{He}/\chi_H)_{HII} \simeq 1$ (this is expected for the early type H II regions chosen for this study), the measured quantity then becomes a measurement of $(\chi_{He} / \chi_H)_{WIM}$. If one further assumes that hydrogen is fully ionized in the WIM, i.e. $(\chi_H)_{WIM} \simeq 1$, then the measured quantity is equal to the fractional ionization of helium in the WIM, or an upper limit on χ_{He} , if χ_H is not 1 in the WIM. The latter assumption is supported by upper limits on the [O I] $\lambda 6300 / \text{H}\alpha$ ratio measured by Reynolds (1989), which lead to the conclusion that $(\chi_H)_{WIM} \geq 0.7$ if $T > 5400$ (the lower limit derived in the same paper), and that $(\chi_H)_{WIM} \geq 0.93$ if the temperature is equal to the generally adopted value of 8000 K. The results for $(\chi_{He} / \chi_H)_{WIM}$ stated below are derived from dividing the $I_{5876} / I_{H\alpha}$ ratio measured for the WIM directions by the average $I_{5876} / I_{H\alpha}$ measured for the H II regions.

Figure 5.1 shows the He I $\lambda 5876$ spectrum for the $l = 194^\circ$, $b = 0^\circ$ direction derived from two 900 second exposures after subtracting the associated “off” directions, which had equal exposure time. An emission line is clearly seen at the same radial velocity found for the $\text{H}\alpha$ emission in this direction. The curve shows a Gaussian fit

to the data, where the mean position was fixed at the location of the $H\alpha$ emission. The width of the Gaussian was also fixed in the fitting process. The width used is calculated from the width of the $H\alpha$ emission feature and the assumption that the gas is at $T = 10^4$. Therefore, the only free parameters in the fit are the baseline parameters and the area of the Gaussian. It is apparent that the mean and width parameters derived from the $H\alpha$ profile are consistent with the He I $\lambda 5876$ emission line. The intensity derived from the Gaussian fit shown in Figure 5.1, when compared with the corresponding $H\alpha$ intensity, leads to $\chi_{He} / \chi_H = 0.60$.

Figure 5.2 shows the He I $\lambda 5876$ spectrum for the $l = 186^\circ$, $b = 0^\circ$ direction. Once again, this spectrum is formed by subtracting an “off” spectrum formed from two 900 second exposures from the spectrum measured “on” the direction of interest. In this case, the emission is clearly fainter than in the $l = 194^\circ$, $b = 0^\circ$ spectrum, even though the $H\alpha$ intensities are the same in the two directions. There appears to be an emission line at the velocity of the $H\alpha$ emission. The Gaussian fit shown was calculated in the same manner as was described for the previous direction, and the resulting position and width of the emission feature are again consistent with the values determined from the $H\alpha$ data. The best fit result for this direction is $\chi_{He} / \chi_H = 0.34$.

The He I $\lambda 5876$ spectrum measured for the $l = 114^\circ$, $b = 0^\circ$ direction is shown in Figure 5.3. For this direction, the baseline is not well determined and thus the results derived from this spectrum are less certain. The baseline irregularity appears to arise from two closely spaced (separated by 20.9 km s^{-1}) atmospheric lines, identified as very weak emission lines of the OH molecule, contaminating this spectrum and the $l = 96^\circ$, $b = 0^\circ$ spectrum. The positions of these lines are indicated by the arrows

in Figure 5.3. When the “off” spectrum is subtracted from the “on”, the variable intensity of these OH lines result in either an excess or a deficiency of emission at their position in Figures 5.3 and 5.4. These atmospheric lines are also present in the $l = 194^\circ$, $b = 0^\circ$ and $l = 186^\circ$, $b = 0^\circ$ directions, but due to the different projection of the earth’s orbital motion onto the lines of sight, they are off to the edge of the spectra in these two cases, and therefore do not affect the He I emission line measurements. For the Gaussian fit shown as the curve in Figure 5.3, the mean and width of the Gaussian was determined from the $H\alpha$ spectrum and was held fixed, while the areas were allowed to vary. Because of the baseline uncertainty, it is possible to fit the spectrum equally well with a curve corresponding to an intensity 35% greater than the curve shown. The fractional ionization of helium with respect to hydrogen calculated from the fit shown in the figure for this direction is $\chi_{He} / \chi_H = 0.37$.

Figure 5.4 shows the He I $\lambda 5876$ spectrum of $l = 96^\circ$, $b = 0^\circ$. The width determined from the $H\alpha$ profile appears narrower than the He I emission line. However, the arrows in Figure 5.4 show the location of the atmospheric OH lines, and it is probably the undersubtraction of these lines that leads to the enhanced emission seen near $v_{lsr} = 20 \text{ km s}^{-1}$. The fit shown leads to $\chi_{He} / \chi_H = 0.59$.

It is instructive to compare the H II region results obtained with the new WHAM spectrometer to the measurements made with the previous PMT based spectrometer (see Reynolds & Tufte 1995). The relative quantum efficiency ratio of the WHAM CCD detector at $\lambda 5876$ relative to $H\alpha$ is approximately 0.70 of the corresponding ratio for the GaAs based PMT detector. The absolute ratios determined for the H II regions appear to be roughly consistent with this.

These results for the $I_{5876} / I_{H\alpha}$ ratio from the new WHAM spectrometer are

somewhat higher than the previous PMT based measurements for the WIM directions (Reynolds & Tufte 1995). The two directions that were measured in both studies are $l = 194^\circ$, $b = 0^\circ$ and $l = 114^\circ$, $b = 0^\circ$. Due to the low signal to noise ratio of the previous data, the data from the two directions were averaged, and a single $\chi_{\text{He}} / \chi_{\text{H}}$ ratio was calculated. The average of the newly measured ratios for the two directions measured in common is approximately 1.7σ higher than the best fit value of the earlier study. The Reynolds & Tufte result is less than 1σ smaller than the average of the values measured here, because of the elevated error bars we are required to use for the new numbers at present. The lower ratios for the $l = 186^\circ$, $b = 0^\circ$ and $l = 114^\circ$, $b = 0^\circ$ directions measured in this study are similar to the earlier result.

5.3 Discussion and Conclusions

The He I $\lambda 5876$ line observations probe the relatively local interstellar medium because extinction by dust limits the region from which emission is observable to within few kpc of the sun. Therefore, the measurement of $\chi_{\text{He}} / \chi_{\text{H}}$ gives direct information on the hardness of the radiation field in the solar neighborhood that is ionizing the observed gas. Specifically,

$$\frac{Q(\text{He}^\circ)}{Q(\text{H}^\circ)} = \left(\frac{\text{He}}{\text{H}} \right) \left(1 + \frac{\text{He}}{\text{H}} \right) \frac{\chi_{\text{He}}}{\chi_{\text{H}}}, \quad (5.2)$$

where $Q(\text{He}^\circ)$ is the flux of helium ionizing photons ($h\nu \geq 24.6$ eV), $Q(\text{H}^\circ)$ is the flux of hydrogen ionizing photons ($h\nu \geq 13.6$ eV), and He / H is the abundance of helium with respect to hydrogen by number. The values of $Q(\text{He}^\circ) / Q(\text{H}^\circ)$ calculated for the WIM directions are shown in Table 5.2. These ratios are significantly lower than would be expected from the radiation field associated with the population of O stars in

the solar neighborhood. Vacca, Garmany, and Shull (1996) have calculated values of $Q(\text{He}^\circ)$ and $Q(\text{H}^\circ)$ as a function of spectral type from O3 to B0.5 considering different luminosity classes separately. They then calculated the expected hydrogen ionizing flux from a sample of 429 O stars within 2.5 kpc of the sun. From their tables, it is also possible to calculate the corresponding helium ionizing flux from the same sample of O stars. The resulting average helium ionizing to hydrogen ionizing photon ratio is

$$\left[\frac{Q(\text{He}^\circ)}{Q(\text{H}^\circ)} \right]_{\text{O stars}} = 0.118. \quad (5.3)$$

The average value calculated from the WIM data presented here is

$$\left[\frac{Q(\text{He}^\circ)}{Q(\text{H}^\circ)} \right]_{\text{WIM}} = 0.052. \quad (5.4)$$

Therefore, the ambient ionizing spectrum resulting from the ensemble of O stars in the solar neighborhood appears to have a significantly harder spectral shape than that which ionizes the WIM.

Other observers using different measurement techniques have reached a similar conclusion. The radio recombination line study of Heiles et al. (1996) probes a different region of the Galaxy, primarily an annular region between a galactocentric radius of 4 kpc and 7 kpc. Their analysis integrated the ionizing radiation emitted over the lifetime of each spectral type, using a standard IMF and constraining the upper mass cutoff for the IMF to be consistent with their measured $\chi_{\text{He}} / \chi_{\text{H}}$. This, combined with a star formation rate set by the known total hydrogen recombination rate, determines the predicted ionizing spectrum. Their measurement of $\chi_{\text{He}} / \chi_{\text{H}} \simeq 0.13$ leads to an upper mass cutoff to the IMF of $M_{\text{u}} \leq 26 M_{\odot}$, which, when combined with the constraint on the total number of hydrogen ionizations, then requires a star

formation rate of $\sim 50 \text{ M}_{\odot} \text{ yr}^{-1}$. This rate is much larger than the star formation rates determined by other means, which are in the range $1.8\text{--}3.6 \text{ M}_{\odot} \text{ yr}^{-1}$. This inconsistency results from their initial assumption that the spectrum of the radiation ionizing the observed gas is the same as the theoretical O star spectrum. Therefore, one possible way to resolve the inconsistency is to assume that the spectrum at the location of the observed gas is different (i.e. softer) than the spectrum emitted by the considered population of O stars.

Observations by Rand (1997) of the He I $\lambda 5876$ line from NGC 891 show that the fractional ionization of helium with respect to hydrogen is also lower than expected in the diffuse ionized gas of this nearby edge-on spiral galaxy. The values of χ_{He} / χ_H obtained by Rand varied from 0.48 to 0.94 with an average value of 0.70. According to Rand, these values are difficult to explain, given the high $[\text{N II}]\lambda 6583 / \text{H}\alpha$ ratios that he measured in the same direction, which require strong heating and thus a hard radiation field.

If O stars are the source of the ionization, these results suggest that their spectrum is modified by processes occurring between the star and the observed emitting gas. A model by Domgörgen and Mathis (1994) explores this scenario. They find that in H II region models with a low excitation parameter, $U \propto \langle n_{\text{photon}} \rangle / \langle n_e \rangle$, the diffuse (and softer) nebular field becomes relatively more important with respect to the direct stellar radiation field, and that as a result the ionization structure of the nebula becomes less dependent on the stellar temperature. More specifically, they construct a composite model of the WIM consisting of 80% matter bounded low density H II region, and 20% ionized H I cloud edge. With a stellar $T_{eff} = 38,000 \text{ K}$ (using stellar atmospheres from Kurucz) they predict $\chi_{He} / \chi_H \approx 0.6$ for the WIM,

which is quite close to the value measured in three of our four background directions. This is particularly impressive since the same model also matches the other unique spectral features of the WIM.

It must also be pointed out that our knowledge of stellar spectra beyond the Lyman limit as a function of spectral type is almost completely theoretical. Observations of emission in this wavelength range are very sparse due to the opacity of the interstellar medium and some of the data that do exist disagree drastically with the models. The B2 II star ϵ CMa was measured by EUVE to emit 30 times more hydrogen ionizing flux than expected by the models (Cassinelli et al. 1995). If this were common among early B stars, then this could explain the softness of the spectrum seen in the WIM, but there is evidence that most B stars do not have such extreme excesses (Kutyrev et al. 1997).

In conclusion, we have clearly detected He I $\lambda 5876$ from the warm ionized medium in three of the four observed directions. In one direction, the $\lambda 5876$ line is clearly fainter relative to $H\alpha$ than in the other directions, and therefore demonstrates that the χ_{He} / χ_H varies significantly with look direction. The average value for χ_{He} / χ_H for the four WIM directions is 0.47 ± 0.17 . This 35% uncertainty is calculated from the worst case difference found in repeat measurements of H II regions and is systematics dominated. The measured χ_{He} / χ_H implies that the radiation field that is ionizing the WIM has a softer spectrum than is emitted by the O stars in the solar neighborhood.

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Nebula	Obs. Date	Spectral Type	$I_{5876}/I_{H\alpha}$
NGC 7000	3-23-96	O6.5 III	0.0171
	7-11-96		0.0169
	7-11-96		0.0199
Rosette	3-10-96	O5.0	0.0120
	3-23-96		0.0164
S184	7-11-96	O5.0	0.0206
S143-S142	7-11-96	O6.0 V + O6.5	0.0183

Table 5.1: O Star Emission Nebula

l	b	χ_{He}/χ_H	$Q(He^{\circ}) / Q(H^{\circ})$
194.0	0.0	0.60	0.066
186.0	0.0	0.34	0.037
114.0	0.0	0.37	0.041
96.0	0.0	0.59	0.065

Table 5.2: χ_{He}/χ_H for WIM directions

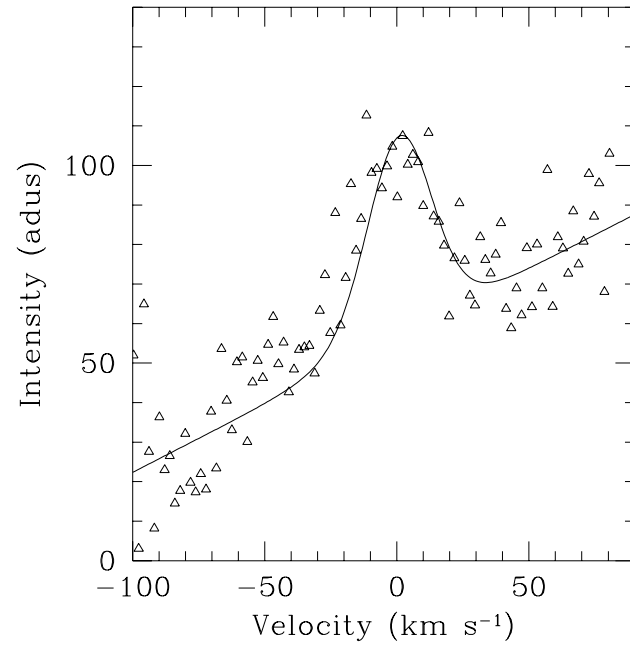


Fig. 5.1.— He I $\lambda 5876$ spectrum of $l = 194^\circ$, $b = 0^\circ$.

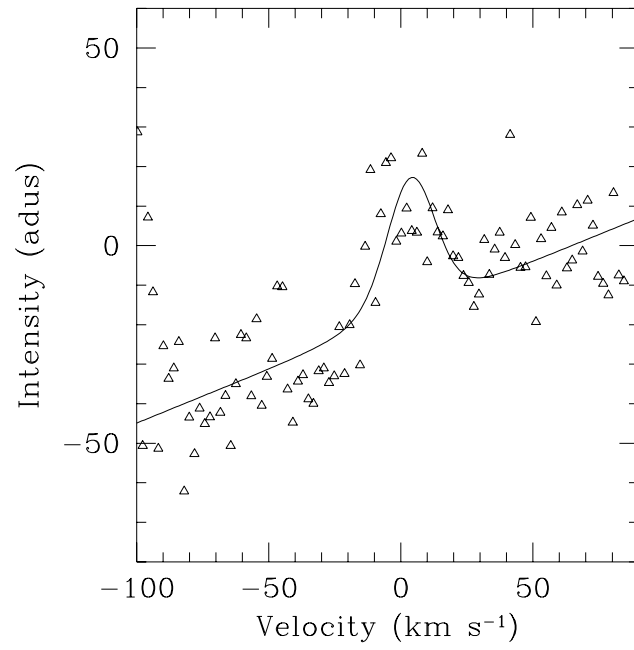


Fig. 5.2.— He I $\lambda 5876$ spectrum of $l = 186^\circ$, $b = 0^\circ$.

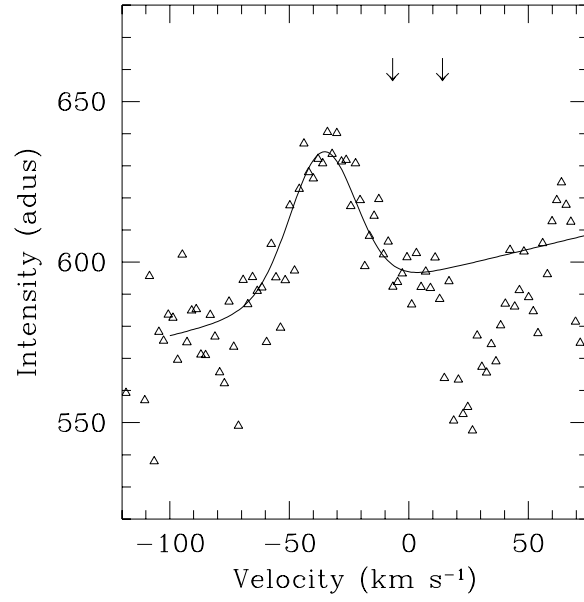


Fig. 5.3.— He I $\lambda 5876$ spectrum of $l = 114^\circ$, $b = 0^\circ$. The arrows show the location of two atmospheric OH lines which contaminate the spectrum.

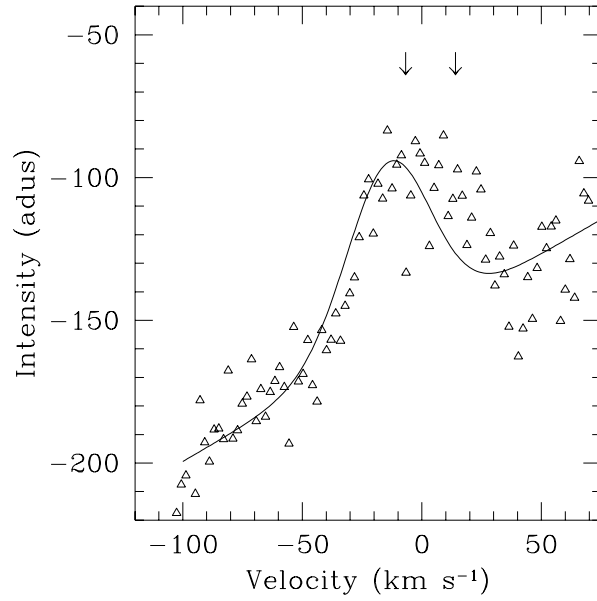


Fig. 5.4.— He I $\lambda 5876$ spectrum of $l = 96^\circ$, $b = 0^\circ$. The arrows show the location of two atmospheric OH lines which contaminate the spectrum.