V1647 ORIONIS: OPTICAL PHOTOMETRIC AND SPECTROSCOPIC MONITORING THROUGH THE 2003–2006 OUTBURST

COLIN ASPIN AND BO REIPURTH

Institute for Astronomy, University of Hawaii, 640 N. A`ohoku Place, Hilo, HI 96720, USA; caa@ifa.hawaii.edu, reipurth@ifa.hawaii.edu

Received 2009 April 14; accepted 2009 June 18; published 2009 September 4

ABSTRACT

We present results from an optical imaging and spectroscopic monitoring campaign on the young, low-mass eruptive variable star V1647 Orionis. The star and associated nebulosity (McNeil’s Nebula) were observed over the period 2004 February to 2006 February with observations commencing a few months after the original outburst event occurred. Using the Gemini North telescope, we obtained multiband optical imaging photometry and medium-resolution long-slit spectroscopy of V1647 Ori on an approximately monthly interval. During this period, V1647 Ori remained at, or close to, peak brightness and then faded by 5 mag to close to its pre-outburst brightness. This implies an outburst timescale of around 27 months. Spectral features seen in both emission and absorption varied considerably during the monitoring period. For example, the H¿ line changed significantly in both intensity and profile. We present and discuss the observed photometric and spectroscopic changes and consider how this eruptive event relates to the early formative stages of low-mass stars.

Key words: accretion, accretion disks – reflection nebulae – stars: individual (V1647 Orionis)

1. INTRODUCTION

When the amateur astronomer Jay McNeil discovered a new nebula in the L1630 molecular cloud in Orion in 2004 January (McNeil 2004), little did he know that it would spark significant and extensive worldwide follow-up investigations. The nebula he discovered has subsequently been designated “McNeil’s Nebula” and its illuminating/exciting star named V1647 Orionis (Samus 2004). To date, around 30 publications have resulted from these studies spanning the electromagnetic spectrum from X-ray to radio wavelengths. In these research papers, considerable speculation has been put forth as to the nature of this eruptive event, yet it is generally agreed that it is the result of a rapid and massive increase in accretion onto the surface of the young star. The reader is referred to the papers by Ābrahām et al. (2004a, 2004b, 2006), Acosta-Pulido et al. (2007), Andrews et al. (2004), Aspin et al. (2006), Aspin et al. (2008, herein ABR08), Aspin et al. (2009a, 2009b), Briceño et al. (2004), Brittain et al. (2007), Fedele et al. (2007a, 2007b), Gibb et al. (2006), Grosso et al. (2005), Kastner et al. (2004, 2006), Kósplá et al. (2005), Kun (2008), McGehee et al. (2004), Mosoni et al. (2005), Muzerolle et al. (2005), Ojha et al. (2005, 2006), Reipurth & Aspin (2004), Retigg et al. (2005), Semkov (2004, 2006), Tsukagoshi et al. (2005), Vacca et al. (2004), Vg et al. (2006), and Walter et al. (2004) for further discussions.

Such eruptive events have been observed in a number of young stars dating back to the observations and discussions by Herbig (1966). In this and subsequent papers, Herbig (1966, 1977, 1989) suggested that these eruptions can be classed as either short-term EXors events (after the progenitor EX Lupi) lasting between a few weeks to a few years, or long-term FUors events (after the progenitor FU Ori) lasting decades to possibly even a century. Much emphasis has been placed on determining what type of outburst has occurred in the case of V1647 Ori, yet more compelling perhaps is understanding whether the FUor and EXor classes are fundamentally similar in their origins or whether they are the result of distinct phenomena occurring on different timescales.

If the outburst of V1647 Ori is a FUor type eruption, then it is the first in almost 40 years since that of V1057 Cyg (Welin 1971). A large-scale effort to monitor V1647 Ori was therefore undertaken involving the use of both Gemini Observatory 8 m telescopes (north and south) and the NASA IRTF 3 m telescope over three observing semesters and utilizing six different facility instruments. This campaign resulted in approximately monthly optical imaging and spectroscopic and near-IR spectroscopic observations of V1647 Ori. In this paper, the first in a series describing these data, we present optical observations spanning the outburst phase, from the first follow-up observations taken soon after its discovery (taken in 2004 February and published in Reipurth & Aspin 2004) to the time when the source had faded to its pre-outburst brightness (in 2006 February). Clearly, the approximately two year lifetime of this event suggests that, at least superficially, the eruption of V1647 Ori is more similar to EXor event than those occurring in FUors.

In this paper, we present optical imaging and spectroscopic observations of V1647 Ori taken between 2004 February and 2006 February and discuss the variations present in these data. The observations and data reduction are described in Section 2. In Section 3, we present details of the temporal changes observed, and in Section 4, we consider what these observations imply regarding the two physical components of the outburst, i.e., the fast wind and the accretion process. Finally, in Section 5 we piece together a timeline for the event and detail the evolution of the eruption as implied by the data from our monitoring campaign.

2. OBSERVATIONS AND DATA REDUCTION

All but one data set presented below were acquired using the “Frederick C. Gillett” Gemini North telescope located on Mauna Kea, Hawaii using the facility optical imager and spectrograph, GMOS-N (Davies et al. 1997; Hook et al. 2004). The imaging observations used standard GMOS-N g′, r′, i′, and z′ filters which were designed to be as close as possible in characteristics to the Sloan Digital Sky Survey (SDSS) filters (Fukugita et al. 1996). For the spectroscopic observations, all but one used the
blue 600 lines mm\(^{-1}\) grating and 0.5 wide long slit. The one exception, the first observation taken, used the red 831 lines mm\(^{-1}\) grating, again with a 0.5 wide long slit. The spectral resolution of data taken were therefore \(\sim 4400\) and \(\sim 1700\), corresponding to 0.34 and 0.45 Å pixel\(^{-1}\), respectively. The complete observation log is presented in Table 1. In total, we acquired photometry at 16 different epochs and spectroscopy at 15 different epochs over the eruption period.

The additional spectrum of V1647 Ori was acquired using the W. M. Keck II telescope located on Mauna Kea, Hawaii, using the facility high-resolution echelle spectrograph HIRES (Vogt et al. 1994). The observations were made on UT 2004 September 24 when the star had an optical \(V\) magnitude of about 17. The spectra covered the wavelength ranges 5800–7150 Å and 7320–8700 Å at a nominal spectral resolution of \(R \sim 46,000\) using a 0.86 wide slit. The on-source exposure time used was 60 minutes.

### 2.1. Imaging Data Reduction and Calibration

All data sets were reduced using the Gemini IRAF data reduction package version 1.8. Specifically, we used the routines gireduce for basic instrument signature removal (trim the images, subtract the master bias image, and divide by the normalized master twilight flat field), and gmosaic for combining the data from the three GMOS-N CCDs into one image. In addition, multiple exposures of the region were co-added using the IRAF routine imcombine. Aperture photometry of V1647 Ori was performed using the Starlink program gaia (Draper et al. 2008). An aperture diameter of 2" was used for the photometric calculation and, due to the bright nebulosity immediately surrounding the star, the sky signal was estimated from numerous apertures located in blank sky regions throughout the image. The full width at half-maximum (FWHM) seeing values in the acquired images are also given in Table 1.

Since some of the data were acquired during nonphotometric conditions, a photometric calibration was achieved relative to a data set taken on a photometric night, UT 2004 February 14. A series of six field stars were used to provide a calibration sequence; their position and magnitudes are listed in Table 2. Since the region containing V1647 Ori is one with active star formation and young stars are well known to be variable, using six field stars as calibrators was considered an acceptable way of minimizing errors introduced by intrinsic variability of the calibrator stars themselves. These stars, together with V1647 Ori, are identified in Figure 1. Using instrumental magnitudes derived from calibrators, we additionally studied the variability of individual calibrators themselves and concluded that uncertainties resulting from source variability were small and within the quoted uncertainties on the V1647 Ori photometry. In addition to this, and to provide a consistency check, we have used the SDSS photometry of V1647 Ori from UT 1998 November 11. These data were part of the early “Orion” release (Finkbeiner et al. 2004) downloaded from Princeton University,\(^1\) and quoted in McGehee et al. (2004). Photometry of the same calibration sequence stars was extracted from the SDSS data and found to be consistent with the photometry from our GMOS images within the associated errors (shown in Table 2).

---

**Notes.**

\(^a\) Modified Julian Date. 2450000+.

\(^b\) SDSS observations taken from the galactic plane Orion release data (Finkbeiner et al. 2004).

\(^c\) GMOS “Sloan” filters.

\(^d\) GMOS Red 831 lines mm\(^{-1}\) grating, blaze wavelength 757 nm, \(R \sim 4396\) with 0.5 slit, simultaneous wavelength coverage 207 nm, 0.034 nm pixel\(^{-1}\).

\(^e\) GMOS Blue 600 lines mm\(^{-1}\) grating, blaze wavelength 461 nm, \(R \sim 1688\) with 0.5 slit, simultaneous wavelength coverage 276 nm, 0.045 nm pixel\(^{-1}\).

\(^f\) Taken at University of Hawaii 2.2 m telescope.

---

**Table 1**

| UT Date (yymmdd) | MJD\(^a\) | Filters and/or Grism Used | Exposure Times (s) | Seeing (\(r’\) Band) |
|-----------------|----------|--------------------------|-------------------|---------------------|
| 981116          | 1134     | \(g,r,i,z^b\)            | \(\ldots\)       | 0.59                |
| 0104214         | 3049     | \(g’,r’,i’,z’^b\), R831\(^d\) | 60,60,60,60       | 0.54                |
| 0104310         | 3074     | \(r’,B600^e\)            | 10,1200           | 0.96                |
| 0104903         | 3251     | \(g’,r’,i’,z’\)          | 30,30,30,30       | 0.53                |
| 0104909         | 3257     | \(B600\)                 | 900               | 0.62                |
| 0104924         | 3272     | HIRES                    | 3600              | ?                   |
| 0104106         | 3284     | \(g’,r’,i’,z’,B600\)     | 30,30,30,30,900   | 0.48                |
| 0104113         | 3322     | \(g’,r’,i’,z’,B600\)     | 30,30,30,1200     | 0.49                |
| 0104121         | 3351     | \(g’,r’,i’,z’,B600\)     | 30,30,30,1200     | 0.58                |
| 0105018         | 3378     | \(g’,r’,i’,z’,B600\)     | 30,30,30,1200     | 0.78                |
| 0105030         | 3612     | \(g’,r’,i’,z’,B600\)     | 30,30,30,30,900   | 0.63                |
| 0105091         | 3634     | \(g’,r’,i’,z’,B600\)     | 30,30,30,30       | 0.48                |
| 0105092         | 3638     | \(B600\)                 | 900               | 0.65                |
| 0105101         | 3650     | \(g’,r’,i’,z’,B600\)     | 30,30,30,30,900   | 0.68                |
| 0105119         | 3692     | \(r’,B600\)              | 30,900            | 0.60                |
| 0105127         | 3701     | \(B600\)                 | 120,3600          | 0.64                |
| 0105128         | 3702     | \(g’,r’,i’,z’,B600\)     | 60,60,60,60,1200  | 0.59                |
| 0105125         | 3720     | \(g’,r’,i’,z’,B600\)     | 60,60,60,60,1200  | 0.49                |
| 0105018         | 3740     | \(g’,r’,i’,z’,B600\)     | 60,60,60,60,1200  | 0.49                |
| 0105030         | 3782     | \(r’,B600\)              | 120,3600          | 0.42                |
| 0105222f        | 4090     | \(R\)                    | 30                | 0.90                |
| 0105021         | 4152     | \(R400\)                 | 3.5 hours         | 0.78                |
| 0105222         | 4153     | \(g’,r’,i’,z’,B600\)     | 600               | 0.60                |

---

**Figure 1.** Region of L1630 including V1647 Ori and McNeil’s Nebula. This \(r’\)-band image was taken on Gemini North using GMOS on UT 2004 February 14. V1647 Ori and the six calibration sequence stars, labeled S1–S6, are identified. North is at the top, east to the left. The scale of the image is 0.144 pixel\(^{-1}\) and the exposure time was 40 s in seeing of FWHM 0.52.

See URL http://photo.astro.princeton.edu/oriondatarelease.
2.2. Spectroscopic Data Reduction

Again, all the Gemini/GMOS data were reduced using the Gemini IRAF data reduction package version 1.8. Specifically, we used the routines gareduce, gtransform, gskysub, gsccrej, and gsextract for basic instrument signature removal (trim the images, subtract the master bias image, and divide by the normalized flat field), wavelength calibration, sky line subtraction, cosmic ray removal, and point-source spectrum extraction, respectively.

The resultant spectra were not flux calibrated nor were atmospheric features removed since we were primarily interested in the spectral features present and their intrinsic variability. In addition, Hα has been shown to be a relatively insensitive diagnostic for accretion rate due to it, and other Balmer lines, becoming optically thick at high accretion rates (Muzerolle et al. 1998), thus negating any requirement for flux determinations.

Figure 8 of Muzerolle et al. (1998) shows model predictions for Hα flux versus accretion rate and it is clear that the relationship saturates at around $10^{-8} \, M_\odot \, yr^{-1}$. Such rather low accretion rates are typically found in “weak-line” T Tauri stars (WTTS) and low-activity “classical” T Tauri stars (CTTS) like DN and DQ Tau and not eruptive variables such as V1647 Ori.

For our HIRES data, a standard reduction was performed, including bias correction, flat-fielding, scattered-light correction, order extraction, and wavelength calibrations using standard routines in the IRAF echelle package.

3. RESULTS

3.1. Optical Photometry

In Figure 2, we show our optical photometry of V1647 Ori from 2004 February to 2006 February for all four passbands, namely $g'$, $r'$, $i'$, and $z'$. The horizontal dashed and dotted lines represent the pre-outburst SDSS brightness of the star (labeled $r_{p0}$). The solid lines simply join the discrete observation points as a guide to the eye. The two long “gaps” in the observation sampling are the 2004 and 2005 summer months when Orion was not visible during the night-time period.

Over the summer of 2004, the source appeared to remain at approximately the same brightness, however, it is likely that some small-scale variability occurred. Our data from the 2004–2005 winter months suggest that variability at the ∼0.5 mag level was present. Such fluctuations were also reported by McGehee et al. (2004), Semkov (2004), Walter et al. (2004), Ojha et al. (2006), Acosta-Pulido et al. (2007), and Fedele et al. (2007b) during several different observing periods.

Over the summer of 2005, V1647 Ori exhibited a small yet significant decline in brightness, approximately the same in all filters, and amounting to ≤1 mag. Immediately after this period, in 2005 September, a major decline phase began. From 2005 September onwards, the trend was for monotonic dimming, with possibly some slowing occurring in late 2005. By the time of our last $g'$-band photometric observation (in early 2006 January), V1647 Ori had already faded by about 4.4 mag from its maximum brightness. At $r'$, we obtained one further observation (in 2006 mid-February) which shows a total fading of over 5 mag. At this time V1647 Ori was already close to its maximum brightness. At $r'$, we obtained one further observation (in 2006 mid-February) which shows a total fading of over 5 mag. At this time V1647 Ori was already close to its maximum brightness.

In late 2006 December we took a short $R$-band exposure of V1647 Ori at the University Hawaii 2.2 m telescope on
Mauna Kea, Hawaii which showed that the source had remained very faint and close to the above pre-outburst SDSS brightness. In addition, ABR08 presented photometry from 2007 February which showed V1647 Ori still close to its pre-outburst brightness ($r' = 23.26 \pm 0.15$).

### 3.1.1. The GMOS and SDSS Photometric Systems

Below, we make use of SDSS photometry and therefore we briefly consider the significance of any differences in photometric systems involved to give confidence to the analysis that follows. As we noted above, the GMOS $g'$, $r'$, $i'$, $z'$ filters were designed to be close to identical to those used by the SDSS survey (Smith et al. 2002). We therefore expect the photometric results from GMOS to be quite similar to those from SDSS. However, to quantify this we have studied the GMOS and SDSS photometry of the calibration sequence used to boot-strap the photometry of V1647 Ori over the monitoring period. In Figure 3, we show the photometry of one representative field star labeled S4 in Figure 1. These photometric values were derived using standard GMOS zeropoints for the observing period (i.e., $ZP(g') = 27.93$, $ZP(r') = 28.18$, $ZP(i') = 27.90$, $ZP(z') = 26.77$; Jorgensen 2009) determined using the observations of Landolt (1992) standard stars observed close in time to the UT 2004 February 14, V1647 Ori observations. The magnitudes of the Landolt standards (in the Johnson–Kron–Cousins photometric system) were transformed to the SDSS photometric system using the relationships derived by Smith et al. (2002). The horizontal lines through the observed data points, while the dashed lines show the magnitude of S4 derived using the GMOS zeropoints for the observing period ($ZP(g') = 27.93$, $ZP(r') = 28.18$, $ZP(i') = 27.90$, $ZP(z') = 26.77$ from Jorgensen 2009) from early 2004. We note that the difference between the two lines is, in all cases, smaller than the associated 1σ error on the data. We conclude, therefore, that the differences between the GMOS and SDSS photometric results from GMOS to be quite similar to those from SDSS. However, to quantify this we have studied the GMOS and SDSS photometry of the calibration sequence used to boot-strap the photometry of V1647 Ori over the monitoring period. In Figure 3, we show the photometry of one representative field star labeled S4 in Figure 1. These photometric values were derived using standard GMOS zeropoints for the observing period (i.e., $ZP(g') = 27.93$, $ZP(r') = 28.18$, $ZP(i') = 27.90$, $ZP(z') = 26.77$; Jorgensen 2009) determined using the observations of Landolt (1992) standard stars observed close in time to the UT 2004 February 14, V1647 Ori observations. The magnitudes of the Landolt standards (in the Johnson–Kron–Cousins photometric system) were transformed to the SDSS photometric system using the relationships derived by Smith et al. (2002). The horizontal lines through the observed data points, while the dashed lines show the magnitude of S4 derived using the GMOS zeropoints for the observing period ($ZP(g') = 27.93$, $ZP(r') = 28.18$, $ZP(i') = 27.90$, $ZP(z') = 26.77$ from Jorgensen 2009) from early 2004. We note that the difference between the two lines is, in all cases, smaller than the associated 1σ error on the data. We conclude, therefore, that the differences between the GMOS and SDSS photometric systems are not significant for the analysis presented below.

#### 3.1.2. Optical Colors

During the whole outburst period, when the source was at or close to its maximum brightness, the optical colors remained approximately constant. In Figure 4, we show the trend in these colors over the monitoring period. The available SDSS pre-outburst colors ($r'–i'$ and $i'–z'$) are also shown as horizontal lines at the right edge of the plot. We note that during the outburst, (1) the optical color indices are larger at shorter wavelengths, (2) both the $r'–i'$ and $i'–z'$ colors are bluer during outburst (by 0.4 and 0.5 mag, respectively), and (3) there is a clear trend for these colors returning to their pre-outburst values as the decline phase progresses. In Figure 5, we show the optical $r'–i'$ and $i'–z'$ colors plotted in a two-color diagram. The colors of V1647 Ori after outburst, but before the decline had started, are the cluster of points near $r'–i' \sim 1.8$ and $i'–z' \sim 1.5$. The point labeled SDSS 1135 is from Modified Julian Date (MJD 2450000+) 1135 which is UT 1998 November 17 (the pre-outburst SDSS source color). The three points labeled 3702, 3729, and 3740 are the source colors at those MJDs and are from the decline phase. These points are significantly closer to the pre-outburst color than those during the outburst. In this plot, we also show the locus of main-sequence dwarf colors (solid line) adapted from Figure 1 of Finlator et al. (2000) and reddening vectors (dashed lines) for a ratio of total to selective absorption $R = A_V/E(B − V) = 3.1$ extending from the extremities of the dwarf locus. These vectors represent a visual extinction $A_V = 5$ mag. We have used the tabular data of D. Finkbeiner (2005, private communication) to
calculate the effective change in $r' - i'$ and $i' - z'$ colors over the $A_V$ range plotted. All the colors of V1647 Ori lie between the two reddening vectors. Dereddening the SDSS 1135 color into the dwarf locus suggests a spectral type of late-K to early-M dwarf. During the outburst phase and through the decline phase, the colors of V1647 Ori first become more blue and then more red. This can be explained by either one, or a combination, of effects. Either a change in intrinsic source color or a color and extinction change has occurred. We explore this effect in a plot of reddening invariant colors versus time shown in Figure 6. Reddening invariant colors are, as the name suggests, colors that do not change as extinction along the line of sight changes. These have been used by McGehee et al. (2004) in their study of V1647 Ori using SDSS photometry and McGehee et al. (2004) in a study of accretion in low-mass young stars. Following Table 4 of McGehee et al. (2004), the reddening invariant color we use, $Q_{riz}$, is defined as

$$Q_{riz} = (r' - i') - 0.987(i' - z')$$

for $R_V = 3.1$. Here, we chose the standard interstellar value of $R_V = 3.1$, however, we note that the change in $Q_{riz}$ for somewhat larger grains, e.g., $R_V = 5.5$, is relatively small ($\sim 10\%$) when compared to the associated uncertainties on the measurements. In Figure 6, we perhaps see a slight trend in $Q_{riz}$ suggesting a larger value as the eruption proceeds into the decline phase. However, the change in value is a small effect with respect to the associated errors. If this difference is real, it would suggest that the intrinsic colors of the source are different out of outburst than in outburst. Such a color change is anticipated due to the added optical continuum luminosity from the enhanced accretion during the outburst (McGehee et al. 2004). Figure 2 of McGehee et al. (2005) plotted numerical model predictions for the $Q_{riz}$ of low-mass young stars versus effective temperature, $T_{eff}$, and as a function of surface gravity, log($g$). P. M. McGehee (2006, private communication) has subsequently produced a version of this plot for us showing the variation in $Q_{riz}$ up to $T_{eff} = 10,000$ K. We do not wish to overinterpret the current data set due to the associated errors, however, comparing the $Q_{riz}$ value with the aforementioned plot, suggests that in its eruptive state the $T_{eff}$ of the optical emission from V1647 Ori determined from its observed $Q_{riz}$ value ($\sim 0.35 \pm 0.15$) is in the range 3000–4000 K. Distinguishing between different surface gravities in this plot would be impossible since the range of $Q_{riz}$ values for log($g$) of 3.5 and 5.5 at this $T_{eff}$ is 0.25–0.35, respectively. Unfortunately, we note that the SDSS pre-eruption value of $Q_{riz}$ from McGehee et al. (2004) is of relatively poor quality ($Q_{riz} = 0.25 \pm 0.23$) due to the faintness of V1647 Ori in $r'$ in 1998 November (MJD 1135, $r' \sim 23.04 \pm 0.22$) and is therefore not considered further.

### 3.1.3. The Pre-outburst to Post-outburst Light Curve

In order to produce as complete a light curve of the 2003–2006 outburst of V1647 Ori as possible, we have combined our photometry with selected published results. In Figure 7, we show our SDSS $i'$-band GMOS photometry from Table 3 together with the $I_C$ photometry of Briceño et al. (2004). We have used the polynomial transformation equation of Ivezić et al. (2007) to estimate SDSS photometric $i'$ values from the Kron–Cousins (Landolt 1983) $I_C$ photometry of Briceño et al. (2004). The transformation equations used were

$$i' = I_C + 0.0307(r' - i')^3 - 0.1163(r' - i')^2 + 0.3341(r' - i') + 0.3584.$$  

Since this (and any) transformation is source color dependent, we assume an intrinsic color for V1647 Ori of $r' - i' = 1.9$, the
et al. reported the \( I_c \) magnitude of \( V1647 \) Ori to be 18.44 ± 0.11 (\( i' = 19.31 \)) in 1999 January, and 20.08 ± 0.5 (\( i' = 20.95 \)) in 1999 December. The data were taken from Acosta-Pulido et al. (2007) and Ojha et al. (2006) although some of the data points in Ojha et al. (2006) were taken from other sources (specifically Reipurth & Aspin 2004; McGehee et al. 2004; Ojha et al. 2005). Unfortunately, there is no \( K \)-band photometry from the spring and early fall of 2006 and therefore we cannot, with certainty, say whether the two light curves are in phase or if there is a relative delay between them. The few points in late 2006 and early 2007 suggest that the major decline phase of \( V1647 \) Ori was somewhat shallower than in the optical although this relies on so few points that it is by no means certain. Nonetheless, the main fact arising from this plot is that the amplitude of the \( K \)-band outburst was considerably less (by 2.2 mag) than in the optical.

Another data set that has good temporal coverage using only one telescope/instrument combination is the \( r' \)-band photometry of Ojha et al. (2006). They present 13 data points covering the period from peak brightness to approximately half way down the steep decline phase. These data are shown in relation to our \( r' \)-band photometry in Figure 8. Due to the extensive overlap of data, we have made no attempt to match the two photometric systems and have merely shifted the Ojha et al. data (filled dots) by +1 mag (fainter) to match our data (open circles) as well as possible. We again note that our final \( r' \)-photometric point (MJD 3782) is very close in value to the pre-outburst (MJD 1135) SDSS \( r' \) photometry value of 23.04 ± 0.22 from McGehee et al. (2004).

We have additionally compiled all published near-IR \( K \)-band (2 \( \mu m \)) photometry and plotted it together with the above \( r' \) light curve in Figure 9. This comparison shows that in the near-IR the change in brightness of \( V1647 \) Ori was less than in the optical and amounted to ∼2.8 mag from the 1998 October Two Micron All Sky Survey (2MASS) photometry (horizontal dashed line). The data were taken from Acosta-Pulido et al. (2007) and Ojha et al. (2006) although some of the data points in Ojha et al. (2006) were taken from other sources (specifically Reipurth & Aspin 2004; McGehee et al. 2004; Ojha et al. 2005). Unfortunately, there is no \( K \)-band photometry from the spring and early fall of 2006 and therefore we cannot, with certainty, say whether the two light curves are in phase or if there is a relative delay between them. The few points in late 2006 and early 2007 suggest that the major decline phase of \( V1647 \) Ori was somewhat shallower at \( K \) than in the optical although this relies on so few points that it is by no means certain. Nonetheless, the main fact arising from this plot is that the amplitude of the \( K \)-band outburst was considerably less (by 2.2 mag) than in the optical.

Figure 7. Optical \( i' \)-band photometry of \( V1647 \) Ori using GMOS (open circles) from 2004 February to 2006 February together with the \( I_c \) photometry from Briceño et al. 2004 (filled circles) from 2003 November to 2004 February, and the Acosta-Pulido et al. (2007) \( I_c \) photometry (filled stars) from 2004 February to 2006 February. The Briceño et al. photometry has been transformed from \( I_c \) to SDSS \( i' \) magnitudes using the transformation equations presented in Ivezic et al. (2007). Additionally, the Briceño et al. data points from MJD 2989 onwards have had an aperture correction applied (+0.62 mag) since the software aperture used by Briceño et al. had a radius of 4 ″ while our aperture radius was 1 ″. The horizontal black dotted line shows the 1998 November SDSS \( i' \) photometric magnitude of the source (\( i' = 20.81 \)). Prior to the time period displayed, Briceño et al. reported the \( I_c \) magnitude of \( V1647 \) Ori to be 18.44 ± 0.11 (\( i' = 19.31 \)) in 1999 January, and 20.08 ± 0.5 (\( i' = 20.95 \)) in 1999 December. The correction used was 0.62 mag. The result of the transformation and aperture correction is to give a reasonably good match (\( \Delta m = 0.14 \)) between our photometry (from MJD 3049) and the closest value from Briceño et al. (from MJD 3036). Also shown in this figure is the \( I_c \) photometry from Acosta-Pulido et al. (2007). We have used the same transformation to SDSS photometry given by Ivezic et al. (2007) and shown in Equation (2) above. In Figure 7, our data are shown as open circles, the Briceño et al. data as filled dots, and the Acosta-Pulido et al. data as filled stars. We additionally note that the pre-outburst (MJD 1135) SDSS \( i' \) photometry from McGehee et al. (2004) (indicated by the black dotted line in Figure 7) is consistent with our last data point (MJD 3782) suggesting that \( V1647 \) Ori had more or less returned to its pre-outburst brightness in 2006 February.

Figure 8. Optical \( r' \)-band photometry of \( V1647 \) Ori using GMOS (open circles) from 2004 February to 2006 February together with the \( R \) photometry from Ojha et al. (2006) (filled stars) from 2004 February to 2005 November. We have shifted the Ojha et al. data vertically by 1 mag to obtain good correspondence between the Gemini and Ojha et al. values throughout the overlap region. The horizontal dotted line is the 1998 November SDSS \( r' \) photometry (\( r' = 23.04 \)).
From Figures 7–9, and previously published data, we can conclude the following.

1. The total duration of the eruptive event was between 847 and 932 days (~28 and 31 months or 2.3–2.6 years). This value assumes that the outburst began between MJD 2850 and 2935 and ended around MJD 3782.

2. The pre-outburst (SDSS, MJD 1135) and post-outburst (Gemini, MJD 3782) brightness of V1647 Ori are very similar.

3. The rise from the pre-outburst brightness to peak brightness took between 145 and 230 days (due to the uncertainty in the date of the start of the eruption, see above).

4. The decline from peak brightness to pre-outburst brightness had two well-defined regions, a shallower phase from late summer of 2004 to the fall of 2005 (∼410 days) when it faded by an additional 3.5 mag.

5. Short-term variability was present during the shallow decline phase. This variability amounted to a peak-to-peak amplitude of ∼0.7 mag.

6. The 2003 eruptive event may have started from a somewhat elevated state. In November 1998 (MJD 1135) the SDSS (r′ and i′) photometry lies between 1 and 2 mag fainter than the brightness when the major eruption phase commenced in late 2003. In addition, our last r′ observation from 2006 February indicates that the star had returned to this fainter level after the outburst had ceased. This behavior is shown graphically in Figure 10 where all of the above optical i′ and Ic data are plotted together with an interpolated i′ magnitude from 2006 February. We note also that the two Ic observations of Briceño et al. (2004) from 1999 suggest that the source possessed significant variability when faint with photometry from 1999 January (Ic = 18.44 ± 0.11 transforms to i′ = 19.31) and 1999 December (Ic = 20.08 ± 0.4 transforms to i′ = 20.95) exhibiting a statistically significant difference of ∼1.6 mag. This may

\[ \text{Using the 2006 February r′ magnitude and the r′–i′ color from our 2006 January observation.} \]
suggest that V1647 Ori underwent a period of significant instability prior to the main eruption phase.

7. The optical (r') and near-IR (K) light curves differ significantly in their amplitude with the optical being around 2.2 mag larger.

8. Some 12 months after the decline to pre-outburst brightness, V1647 Ori remained optically very faint and close to its pre-outburst brightness.

3.2. Optical Spectroscopic Features and Their Evolution

In this section, we discuss our optical spectroscopic observations over the monitoring period but defer discussion of the strongest feature in the spectra, the Hα line, to Section 3.3. However, we do consider the Hα line when describing the HIRES echelle spectra below as an introduction to the line properties. We note that all radial velocities quoted below have been corrected to be heliocentric and that to convert to velocities in the stars frame of reference the reader should use the rest velocity of the L1630 molecular cloud measured directly on V1647 Ori by Andrews et al. (2004) to be +10 km s⁻¹ with respect to the L1630 molecular cloud measured directly on V1647 Ori stars frame of reference the reader should use the rest velocity corrected to be heliocentric and that to convert to velocities in the

We show a GMOS spectrum of V1647 Ori in Figure 11 selected as representative of the 15 GMOS spectra we obtained between 2004 February and 2006 February. These are the data from 2004 October 6 since it has good signal-to-noise ratio (S/N) and shows most of the spectral features present. The source shows a rising red continuum with strong Hα emission exhibiting a P Cygni profile. There are numerous weak features in the spectrum including the Na D absorption at 5890 and 5896 Å. Table 4 lists the lines found in the spectra and their variability over the monitoring period. We have not measured equivalent widths for the lines since, as we shall discuss below, such a value is affected significantly by the highly variable optical continuum flux observed. Instead, we have merely specified whether the lines are in absorption or emission, or not present. As well as permitted emission lines of Fe, O, and Mg, numerous forbidden emission features, i.e., [Fe II], [O I], [S II], and [Ca II], are observed. In Figures 12–14, 16, and 17, we show extracted wavelength ranges around the Na D absorption lines, and the [S II], [O I] (6300 Å), Fe (6400–6550 Å), and the Ca emission lines (7200–7400 Å), respectively.

3.2.1. Na D Absorption Lines

The Na D neutral absorption line profiles are shown in Figure 12. The first observation of the monitoring period, taken on 2004 February 14, only showed weak Na D absorption. This was blueshifted by ~200 km s⁻¹ with respect to the rest wavelengths of the lines. Between 2004 February 14 and March 10, the Na line profiles change significantly, they became stronger and considerably broader. On 2004 March 10, they were blueshifted by ~280 km s⁻¹ with respect to the line rest wavelengths. Also at this time two emission features, presumably Na D lines, are present at 5895.3 Å (+270 km s⁻¹ with respect to the bluer Na D line rest wavelength of 5889.9 Å) and 5899.7 Å (+190 km s⁻¹ with respect to the redder Na D line rest wavelength of 5895.9 Å). These appear to be real features since emission is also seen in the lower resolution 2004 February 18 spectrum from Briceño et al. (2004). An absorption feature is seen at 5877 Å and its depth appears correlated to the strength of the Na D lines. This absorption feature is similar in width (~160 km s⁻¹) to the main Na I lines and could possibly be a highly blueshifted component of Na absorption appearing at ~640 km s⁻¹ with respect to the shorter wavelength Na I line at 5890 Å. An alternative explanation for the presence of this line is that it is the He I line at 5876 Å in absorption. Perhaps this is the more reasonable interpretation and is supported by the fact that we also observe He I lines at 6678 and 7766 Å in absorption.

After the summer months of 2004, the first spectrum taken was on 2004 September 9 and showed the Na D lines even stronger than on March 10. Between September 9 and 2005 January 8 (the last observation before the 2005 summer break), the Na lines vary in intensity, becoming weaker and stronger in successive months. The Na lines appear strongest on 2005 January 8 and are blueshifted with respect to the rest wavelength by ~−150 km s⁻¹. Large changes occur between 2005 January 8 and the next observation on 2005 August 30 where the Na lines have weakened significantly and appear less blueshifted (~−100 km s⁻¹). In the UT 2005 September 25 spectrum, the Na lines possibly appear weakly in emission and very close to their rest wavelengths (~40 km s⁻¹). From 2005 October 13 to the end of the monitoring period, the Na lines have more or less disappeared.

Highly structured Na I 5890 and 5895 Å lines are not uncommon among eruptive variables, specifically FUors. Examples include FU Ori itself (Bastian & Mundt 1985; Hartmann & Calvet 1995), BBW 76 (Reipurth et al. 2002), V1057 Cyg (Bastian & Mundt 1985; Herbig et al. 2003), and V1515 Cyg (Bastian & Mundt 1985). It was noted by Bastian & Mundt (1985) that all Na D profiles appear very similar to each other.

4 Kindly made available to us by C. Briceño.
Figure 11. Optical GMOS spectrum of V1647 Ori from 2004 October 6. This is a representative spectrum of the 15 acquired between 2004 February and 2006 February. In 2004 October, V1647 Ori had a magnitude of $r' = 18.03$ and was in its shallow decline phase. All identifiable absorption and emission lines are marked including the telluric O$_2$ and O$_2$-A bands. Note the prevalence of weak permitted and forbidden Fe emission lines. The major intrinsic feature is the strong H$_\alpha$ emission line with a blueshifted absorption component forming a P Cygni type profile. Table 4 shows the variation in spectral features over the monitoring period.

Table 4
Identification of Spectral Features

| MJD  | YY   | MM   | DD   | [Fe ii] 5226 | He i 5876 | Na D 5890+5896 | [O i] 6300 | Si ii 6347 | [Fe ii] 6432 | Fe i 6495 | Fe ii 6517 | H$_\alpha$ 6563 | [O ii] 6713+6731 |
|------|------|------|------|-------------|-----------|----------------|-----------|---------|-------------|---------|---------|-------------|---------------|
|      | 04   | 02   | 14   | –           | –         | A              | –         | A       | –           | –       | –       | –           | –             |
| 3049 | 04   | 03   | 18   | E           | WE        | A              | E         | WE      | –           | WE      | –       | –           | –             |
| 3053 | 04   | 09   | 10   | E           | WE        | WA             | A         | WA      | –           | –       | –       | –           | –             |
| 3074 | 04   | 10   | 03   | –           | –         | AE             | AE        | A       | A           | –       | ?       | –           | –             |
| 3251 | 04   | 11   | 06   | –           | –         | A              | A         | A       | A           | –       | –       | –           | –             |
| 3257 | 04   | 12   | 12   | –           | –         | –              | E         | WE      | –           | WE      | –       | –           | –             |
| 3284 | 04   | 08   | 08   | –           | –         | –              | E         | WE      | –           | WE      | –       | –           | –             |
| 3322 | 05   | 30   | 25   | –           | –         | –              | –         | –       | –           | –       | –       | –           | –             |
| 3378 | 05   | 19   | 13   | –           | –         | –              | –         | –       | –           | –       | –       | –           | –             |
| 3612 | 05   | 27   | 19   | –           | –         | –              | –         | –       | –           | –       | –       | –           | –             |
| 3634 | 05   | 05   | 13   | –           | –         | –              | –         | –       | –           | –       | –       | –           | –             |
| 3656 | 05   | 06   | 12   | –           | –         | –              | –         | –       | –           | –       | –       | –           | –             |
| 3692 | 05   | 06   | 11   | –           | –         | –              | –         | –       | –           | –       | –       | –           | –             |
| 3701 | 05   | 06   | 10   | –           | –         | –              | –         | –       | –           | –       | –       | –           | –             |
| 3702 | 05   | 07   | 09   | –           | –         | –              | –         | –       | –           | –       | –       | –           | –             |
| 3740 | 06   | 06   | 05   | –           | –         | –              | –         | –       | –           | –       | –       | –           | –             |
| 3782 | 06   | 07   | 06   | –           | –         | –              | –         | –       | –           | –       | –       | –           | –             |
| 4152 | 07   | 07   | 07   | –           | –         | –              | –         | –       | –           | –       | –       | –           | –             |

Notes.

a E, emission; A, absorption; WE, weak; P, P Cygni profile; X, bad pixels at location of line; ?, inconclusive; –, no line evident; O, outside observed spectral range.

This also is the case if we compare these three FUors to the profile of BBW 76 from Reipurth et al. (2002). The time-series of BBW 76 by Reipurth et al. (2002) is particularly interesting since they observed multiple minima in both Na lines which were predominantly blueshifted by up to $-300$ km s$^{-1}$. In addition, between 1985 and 1994 the lines varied significantly in both width and blueshifted velocity. We note here that the generally accepted interpretation of such Na i line structure is that an intense stellar wind, resulting from enhanced accretion, forms shell-like expanding/outflow structures which are accelerated close to the star/disk and slow with increasing distance (Bastian & Mundt 1985; Reipurth et al. 2002). Since our data do not have the spectral resolution of the BBW 76 data ($R \sim 2000$ as opposed to 20,000 for BBW 76) we cannot resolve individual shell components. However, the large line width and the blueshifted nature of the absorption implies in itself that we are seeing considerable outflowing material with a significant velocity gradient. If the absorption feature at 5877 Å is indeed a high-velocity
blueshifted Na component, then within one month (2004 February 14 to March 10) the wind dramatically increases in density (line depth) and velocity (line width/range) with a maximum of over $-600$ km s$^{-1}$. For a spherically symmetric wind, the fact that we do not observe a symmetric profile, one with both blueshifted and redshifted absorption, suggests that we only have an unobscured line of sight toward the absorbing material that is expanding toward us. This is likely the result of obscuration by the circumstellar/accretion disk and implies that the material lies close to the star and is accelerated, in this location, to high velocities (see the discussion of Bastian & Mundt 1985 for more details). This is consistent with the inferred inclination of the axis of the nebula (McNeil’s Nebula) to the line of sight of $\sim 60^\circ$ (Acosta-Pulido et al. 2007; ABR08), since we are looking onto the star/disk through the nebula.

3.2.2. [S ii] Emission Lines

Fedele et al. (2007b) and ABR08 found [S ii] emission at 6717 and 6731 Å in the spectrum of V1647 Ori in 2006 January and 2007 February, respectively. In Figure 13, we trace the evolution of this emission from 2004 February to 2007 February. We see that there are weak [S ii] lines present as early as 2004 March 10 and they are seen in all our spectra up until the end of the monitoring period. The lines fall at their rest wavelength with the resolution we have, and vary in both intensity and ratio with time. The ratio $6717/6731$ is greater than unity from 2004 March 10 to 2006 January 5, and then, in our last two spectra, drops to less than unity. From 2005 October 13 onwards, the lines appear to be double-peaked with a separation of the peaks being $\sim 40$ km s$^{-1}$. However, our spectral resolution is insufficient to really study this structure in any detail or even to be certain of its existence. If it is present then perhaps we are detecting red and blueshifted shocks from the star.

We have not derived electron densities, $n_e$, from the ratio of the two sulfur lines (6717/6731, the ratio is only weakly dependent on temperature) from all our spectra since the S/N of the line detection is, in many cases, insufficient. However, for five observations, 2005 October 13, 2005 November 27, 2005 December 25, 2006 January 5, and 2007 February 21, we consider the lines well enough detected. The derived ratios for $6717/6731$ are $1.3 \pm 0.1, 1.4 \pm 0.1, 1.4 \pm 0.1, 1.4 \pm 0.1$, and $0.8 \pm 0.1$ which give values of $n_e$, derived using the IRAF nebular.temden program for a temperature of $10^4$ K, of $26-240, <121, <121, <121, and 910-2230$ cm$^{-3}$, respectively. The range of values and the upper limit arise from the uncertainties on the line ratios and the failure of temden at ratios $>1.4$, respectively. It seems, therefore, that the forbidden sulfur lines are formed in a rarefied gas which becomes denser toward the end of our monitoring period although, even then, it is far from the [S ii] critical density of 20,000 cm$^{-3}$.

3.2.3. [O i] 6300 Å Line

The [O i] emission line at 6300 Å varied significantly in intensity over the monitoring period. The changes are shown in Figure 14. We note that the wavelength of this line is coincident with a strong night-sky emission line although we consider sky subtraction to be generally very good. At the start of the observing period the line was absent and appeared some time between 2004 February 14 and 2004 September 9 (unfortunately the spectrum from 2004 March 10 had a series of bad pixels at this wavelength and it was not possible to determine the presence of the line). From 2004 September 9 onwards, the [O i] line appeared to have an extended blue wing which persisted until the line faded. The blue wing extended to $\sim -270$ km s$^{-1}$ from the line rest wavelength and Figure 15 shows its profile from the spectrum taken on UT 2004 November 13. The line continued to increase in intensity to 2005 January 8 where it was at a maximum. Subsequently, the line faded but was still perhaps detectable in the low-resolution spectrum taken on 2007 February 21. We note that the spectrum from 2006 February 16.

Figure 12. Variation in Na D 5890 and 5896 Å absorption profiles over the monitoring period. The continuum has been subtracted in these spectra and they have not been smoothed so as to preserve the intrinsic noise structure for comparative purposes. The vertical dashed lines show the Na lines rest wavelengths. In 2004 February, the Na lines were barely visible but soon became very pronounced and blue-shifted with respect to the rest wavelength. The lines are highly variable in depth and shape until 2005 August (at the beginning of the steep decline phase) when they had faded considerably. For the remainder of the monitoring period (2005 September to 2007 February) the lines were barely present.
Figure 13. Variation in the shock-excited [S\textsc{ii}] 6717 and 6731 Å emission line intensity and ratios over the monitoring period. The vertical dashed lines show the [S\textsc{ii}] lines rest wavelengths. In 2004 February, the [S\textsc{ii}] lines were not present. In the spectrum taken on 2004 March 10 however, they are seen faintly. The intensity and ratio of the lines varies from then to the end of the monitoring period. Again, the spectra have been continuum subtracted and not smoothed.

Figure 14. Variation in the [O\textsc{i}] 6300 Å emission line intensity over the monitoring period. The vertical dashed line shows the [O\textsc{i}] line rest wavelength. The line is not present on 2004 February 14 but is strong in the spectrum taken on UT 2004 September 9. The spectrum from 2004 March 10 had considerable noise spikes at the wavelength of [O\textsc{i}] which have been removed by interpolation. The line intensity mostly increases from UT 2004 September 9 to UT 2005 January 8 when it begins to decline. The spectrum from 2006 January 5 had some noise spikes (which have been removed) but the wings of the [O\textsc{i}] line can still be seen. The line is perhaps just detected in the UT 2007 February spectrum. Again, the spectra have been continuum subtracted and not smoothed. The [O\textsc{i}] emission line is asymmetric in shape with an extended blueshifted wing. We note the presence of an unidentified absorption feature at \(\sim6350\) Å.

had bad pixels at 6300 Å which were removed by interpolation. However, the wings of the [O\textsc{i}] line are still detected.

The presence of forbidden emission lines in the optical spectra of young CTTSs has been studied by several authors (e.g., Edwards et al. 1987; Hartigan et al. 1995) and is generally associated with low-density outflowing gas. Typically, only blueshifted emission wings are observed which has led to the model that the circumstellar disk occults redshifted emission from a latitude-dependent wind with the highest gas velocities at the polar region of the star. From the models of Edwards et al. (1987), the only way to obtain [O\textsc{i}] profiles like those seen in V1647 Ori, i.e., single-peaked with an extended blueshifted wing, is to view the wind from a low inclination angle, specifically, \(i < 45^\circ\) for a wind opening angle \(\theta < 30^\circ\) (see Figure 10 in Edwards et al. 1987). This upper limit to the inclination angle is somewhat at odds with
In the wavelength range 6400–6540 Å, there are four Fe emission lines from permitted (three) and forbidden (one) transitions. These are [Fe II] at 6432 Å, Fe II at 6457 Å, Fe I at 6495 Å, and Fe II at 6517 Å. The evolution of these lines is shown in Figure 16. The [Fe II] 6432 Å line is seen from the start of the monitoring period, 2004 February 14 to 2005 November 27 with peak emission occurring on 2005 January 8. The Fe II 6457 Å line is relatively weak in all spectra from 2004 September 9 to 2005 November 27 and peak emission also occurs on 2005 January 8. The Fe II 6495 Å line is seen from 2004 March 10 to 2005 October 13 and again, the emission peaks on 2005 January 8. Finally, the Fe II 6517 Å line is visible from 2004 February 14 to 2005 November 19 and it is strongest on 2005 January 8. The Fe II 6495 Å line is consistently broader than either the Fe II or [Fe II] lines and it has a FWHM of 250 km s\(^{-1}\) compared to 150 km s\(^{-1}\) for the permitted and forbidden ionized Fe lines. For comparison, the [Ca II] line at 7292 Å has a FWHM of 150 km s\(^{-1}\) while a CuAr arc lamp line located near the [Ca II] line has a FWHM of 130 km s\(^{-1}\).

In Figure 17, we also identify an additional [Fe II] emission line at 7388 Å. Its behavior is similar to that of the [Fe II] 6517 Å line although it disappears slightly earlier (2005 September 25 as opposed to 2005 November 27). It also peaks in intensity on 2005 January 8.

The behavior of the [Fe II] line at 6432 Å seems consistent with that of the [Si II] lines described above and suggests a common excitation mechanism such as shock-excitation. The permitted Fe lines are generally considered as indicators of chromospheric activity in T Tauri stars (Beristain et al. 1998) and therefore should show similar behavior to the [Ca II] lines described below.

### 3.2.5. [Ca II] Emission Lines

We have identified two [Ca II] emission lines at 7292 and 7324 Å. The variations in these lines with time are plotted in Figure 17. Our 2004 February 14 and March 10 spectra did not include these lines, however, they are present in the spectra from 2004 September 9 to 2007 February 21. The [Ca II] lines are well detected from 2004 September 9 to 2005 August 30 and are weakly detected in the 2005 September 25 to 2005 October 13 spectra. They are absent in the spectra taken after the latter date. The [Ca II] lines are unresolved with respect to the line width defined by the CuAr arc lamp lines (see above).

### 3.2.6. The Keck II HIRES Spectrum

The HIRES echelle spectra of V1647 Ori shows a number of both absorption and emission features with high-resolution providing information on their structure. Below, we consider each spectral feature and discuss their appearance.

The H\(\alpha\) line at high-spectral resolution is particularly interesting. It has been previously observed at low-resolution to exhibit a P Cygni profile with a strong blueshifted absorption and very broad emission (e.g., Reipurth & Aspin 2004; Walter et al. 2004). Our HIRES spectra are the highest spectral resolution data (\(R \sim 46,000\)) obtained on the source to date. Figure 18 shows the profile of H\(\alpha\) from 2004 September 24, some 10 months after the eruption occurred. At this time the source had a magnitude of \(r' \sim 18\). The emission feature is asymmetric with an extended red wing reaching around +360 km s\(^{-1}\). The blueshifted absorption seems to begin close to the line rest wavelength and continues out to \(\sim -360\) km s\(^{-1}\), the slight peak at this velocity appears to be where the blue wing of the emission line reaches the continuum level. This would make the emission symmetric about the line rest wavelength with a full width zero intensity (FWZI) of around 720 km s\(^{-1}\). The emission red wing appears to exponentially decay to the continuum. It is difficult to say where the line emission peak flux is located due to the overlying absorption. The absorption component appears to start close to the line rest wavelength and creates the very asymmetric profile seen in the blue wing. This is probably formed by a combination of two absorption features. The blueshifted absorption goes below the local continuum level and is flat-bottomed from about \(-100\) to \(-330\) km s\(^{-1}\). In Figure 18, we show the H\(\alpha\) profile together with a best-fit profile (top) fitting the red wing of the emission only. Also shown (bottom) is the residual profile showing the difference between the observed and best-fit profile. The profile of the fitted feature is Lorentzian (either a Gaussian or Voigt profile fit the red wing well). We note that the profile of an unresolved arc line (ThAr) is best-fit with a Gaussian of FWHM \(\sim 6\) km s\(^{-1}\). The residual profile shown in Figure 18 exhibits extensive absorption superimposed on the H\(\alpha\) emission and clearly shows two distinct components, one at the H\(\alpha\) rest wavelength (the dot-dashed line) with a FWHM of \(\sim 50\) km s\(^{-1}\), and the other extending from the continuum at \(-70\) km s\(^{-1}\) out to \(-350\) km s\(^{-1}\). The more highly blueshifted component has a more complex profile with the absorption being deepest at around \(-100\) km s\(^{-1}\) then decreasing in an asymptotic manner to \(-290\) km s\(^{-1}\) where the decrease becomes linear to the level of the continuum.
Fe lines in the 6400–6540 Å region of the spectra. These lines are \[\text{Fe II}\] at 6432 Å, \[\text{Fe II}\] at 6457 Å, \[\text{Fe I}\] at 6495 Å, and \[\text{Fe II}\] at 6517 Å. The vertical dot-dashed lines are at the line rest wavelengths. The lines vary in intensity together, getting stronger from UT 2004 February 14 (when they are very weak) to UT 2005 January 8. After this, they decline in intensity until UT 2005 November 27 when they are not detected. Again, the spectra have been continuum subtracted and not smoothed.

\[\text{Ca II}\] lines in the 7200–7400 Å region of the spectra. Specifically, at 7261, 7292, and 7324 Å. The vertical dot-dashed lines are at the line rest wavelengths. At the long-wavelength edge of the plot is an \[\text{Fe II}\] at 7388 Å. Our UT 2004 February 14 and UT 2004 March 10 spectra did not cover this region of the spectra, however, from UT 2004 September 9 to UT 2005 October 13, the \[\text{Ca II}\] lines are visible. After this they are not present. Again, the spectra have been continuum subtracted and not smoothed.

We detect the K\textsc{i} 7665 and 7699 Å lines which also possesses P Cygni profiles. The shorter wavelength line is located in the middle of the atmospheric O\textsc{II} absorption band, however, the P Cygni structure is still evident and consistent with the longer wavelength line. The P Cygni absorption component has a velocity offset from the rest wavelength of $-160$ km s$^{-1}$. The 2004 October low-resolution GMOS spectrum shows these K\textsc{i} lines which also display P Cygni profiles.

The Na D lines are present in absorption only and are broad, relatively symmetric, and blueshifted by about $-145$ km s$^{-1}$. These are consistent with our low-resolution GMOS spectrum from our nearest date, UT 2004 October 6. The S/N in the HIRES spectrum at the wavelength of the Na D lines is insufficient (S/N pixel$^{-1}$ $\sim$5.5) to say if a narrow interstellar Na D absorption component is present.

The 6347 and 6371 Å Si\textsc{II} lines are seen in absorption and both are shallow and broad (FWHM$\sim$130 km s$^{-1}$). These lines are also seen in our low-resolution GMOS spectra from 2004 October 6.

The Ca\textsc{II} line at 8498 and 8662 Å lines are strongly in emission (their equivalent width are 10 and 8 Å, respectively). The middle of the Ca\textsc{II} triple lines, at 8542 Å, is unfortunately
spectrum confirming their identification in the HIRES data. The emission. This is also seen in our 2005 October 6 low-resolution detection in the HIRES data.

Also they are relatively narrow with respect to other emission lines with a FWHM of 12 km s\(^{-1}\). Below is the difference between the observed and model profile. There appears to be significant H\(\alpha\) absorption at zero velocity together with blueshifted absorption extending from \(-80\) to \(-360\) km s\(^{-1}\). The minimum in the H\(\alpha\) absorption lies at 120 km s\(^{-1}\).

Strong line absorption at 7773 Å corresponds to the O\(i\) triplet lines. This line is at the end of an echelle order and so no further information is obtainable. In our low-resolution GMOS spectra, the O\(i\) triplet is highly blended but is seen in absorption from 2004 September 3 to 2005 October 13. After this the lines are either not present or hidden in the noise.

We detect two very weak Mg\(\text{ii}\) absorption lines at 7877 and 7897 Å. In our low-resolution spectrum from 2005 October 6, these lines are seen weakly in absorption confirming their detection in the HIRES data.

It seems that the 6300 and 6363 Å O\(i\) lines are very weak in emission. This is also seen in our 2005 October 6 low-resolution spectrum confirming their identification in the HIRES data. The S/N of the HIRES detection is low but the lines appear real. Also they are relatively narrow with respect to other emission lines with a FWHM of \(-40\) km s\(^{-1}\).

A number of broad, weak emission lines of Fe\(i\) (six lines) and Ti\(i\) (two lines) are present. The FWHM of both the Fe\(i\) and Ti\(i\) line is \(-100\) km s\(^{-1}\). There is a hint that the lines are double-peaked but, with the S/N present, it is difficult to say if this is correct or if they are broadened single lines or possibly overlapping pairs.

From the above detailed description of the HIRES spectra, we can conclude that the features present agree well with those seen in the 2004 October 6 low-resolution GMOS spectrum. The HIRES spectrum also confirms that some of the weaker lines seen in the GMOS spectrum (with low S/N) are real.

### 3.3. The Variability of the H\(\alpha\) Line

In Figure 20, we present the H\(\alpha\) profile of V1647 Ori over the period 2004 February to 2007 February. These 15 observations show considerable variation in both emission and absorption characteristics as the eruption proceeded and during the decline phase. The bottom right tile of the plot shows the instrumental profile (CuAr arc lamp lines) close to the wavelength of H\(\alpha\). Figure 21 shows an expanded view of the H\(\alpha\) profiles with an x-axis of velocity offset in km s\(^{-1}\) from the rest wavelength of H\(\alpha\) (6562.8 Å).

Qualitatively, the H\(\alpha\) profile varies significantly even between consecutive observations. H\(\alpha\) typically exhibits P Cygni type structure with an emission component and a blueshifted absorption feature. Throughout the outburst period, the emission component is extremely broad with a maximum FWZI of close to 1500 km s\(^{-1}\), however, it decreases in width as the eruption progresses. The width of the blueshifted absorption is over 600 km s\(^{-1}\) at the start of the monitoring period and also becomes smaller with time. Additionally, the absorption shifts toward the rest wavelength of H\(\alpha\) as well as decreasing in strength. Throughout the first 18 months of the monitoring period, the absorption component absorbs below the continuum (see panels 2004FEB14 to 2005AUG30 in Figures 20 and 21). After 2005AUG30, the wind absorption appears to be not strong enough to absorb below the continuum level and merely eats into the H\(\alpha\) emission. Finally, we note that the H\(\alpha\) emission and absorption components are not of sufficient S/N in our last three data sets (2005DEC25, 2006JAN05, and 2006FEB16) to reliably determine the extent of the absorption on the emission component.

We quantify the H\(\alpha\) profile structure and variations in both Table 5 and Figure 22. In the latter, we show several measures

---

**Figure 18.** H\(\alpha\) line profile from the HIRES spectrum of V1647 Ori. At the top is the observed spectrum overlaid with a model Gaussian profile that fits the red wing of the line. Note the wings of the profile extend to \(\pm 360\) km s\(^{-1}\) from the nominal wavelength of H\(\alpha\) (dot-dashed line). For comparison, an arc lamp line lying close to the wavelength of H\(\alpha\) has a FWHM Gaussian profile of 12 km s\(^{-1}\). Below is the difference between the observed and model profile.

**Figure 19.** HIRES spectrum of V1647 Ori containing the Ca\(\text{ii}\) line at 8662 Å. Note the triangular profile of the line. Also present are two faint Fe\(i\) lines which seem to show a double-peaked nature.
of Hα profile structure, specifically the equivalent width of the emission component ($W_{\lambda}$, top-right), the full width zero intensity ($I_{\text{max}}$, FWZI) and the full width at $I_{\text{max}}/40$ (2.5%) intensity (FW2.5%, middle-left), the velocity width of the blueshifted absorption feature ($\Delta V$, middle-right), the velocity offset of the wavelength of deepest or “characteristic” absorption ($V_{\text{char}}$, bottom-left), and the wavelength of peak (i.e., $I_{\text{max}}$) Hα emission ($\lambda_{\text{peak}}$, bottom-right). We show the FWZI and FW2.5% variations since it is interesting to both compare the behavior of these measures of profile width and directly relate the latter to values seen in other stars, specifically from Reipurth et al. (1996, henceforth RPL96). RPL96 pointed out that the FW2.5% is considerably less sensitive to noise than FWZI.

From Figure 22 we see that as the eruption proceeded from early 2004 to early 2006 (detailed in the r′-band brightness variations shown in the top-left panel), the behavior of the Hα profile measures was rather complex. Clearly, detailed numerical modeling of the profile changes is required to characterize and quantify the physical and geometric properties of the accretion and outflow during the eruption (see, for example, Kurosawa et al. 2006). However, we consider that the above five empirically defined quantities may give us some useful insight into the nature of the variability observed.

3.3.1. Hα Equivalent Width in the “High Plateau” Phase

Our measurements of the variation of the Hα emission $W_{\text{H}\alpha}$ versus time are shown as the filled circles in the top-right panel of Figure 22. This demonstrates that from soon after the outburst began (2004 February) through to the start of the major decline phase (2005 September), $W_{\text{H}\alpha}$ was approximately constant, with perhaps a slight decline, and had a mean value around $-30 \text{ Å}$. This time period was referred to as the “high plateau” phase of the outburst decline by Acosta-Pulido et al. (2007) and we herein adopt this descriptive designation. Over the high plateau period, the brightness of V1647 Ori showed a decline of about $\sim 1.2$ mag (see the top-left panel in Figure 22). This would result in an increase in $W_{\lambda}$ for constant Hα emission flux. Our average value of $W_{\text{H}\alpha} \sim -30 \text{ Å}$ is consistent with the values presented in both Walter et al. (2004) and Ojha et al. (2006) whose data are also shown in Figure 22 (as open squares and open circles, respectively). There are clearly short-term fluctuations in $W_{\text{H}\alpha}$ in all three data sets. These variations were considered possibly periodic in nature by Walter et al. (2004) and Ojha et al. (2006). Acosta-Pulido et al. (2007) detected short-term optical photometric variability in their light curves and, from a Fourier periodicity search, derived a 56 day period. They commented that one possible interpretation of such a periodicity would be variable obscuration by circumstellar disk material orbiting the young star.

3.3.2. Timescales of Hα Variability

If such a periodicity is present in broadband optical photometry or $W_{\text{H}\alpha}$, or both, then it is clearly an important phenomenon for investigating the physics occurring during the V1647 Ori outburst. In our spectroscopic data, we perhaps see some temporal structure that could be interpreted as periodic in nature.
For example, in the period 2004 September to 2005 January, the variations appear reasonably well phased with the data of Walter et al. (open squares). However, with so few observations (five) and their temporal spacing, we cannot confirm that such periodic variability exists. What we can do is characterize the shortest timescale upon which variability is found. To do this, we consider the minimum separation between consecutive observations (combining all data points) which show statistically significant variability and we adopt a > 10% change in $W_{\text{H}\alpha}$ as indicating statistical significance. There exists in the data only one date (MJD 3378) when observations were taken less than 1 day of each other (i.e., on the same night). Upon this date, the $W_{\lambda}$ values are consistent to better than 10%. Next, there are three occurrences of consecutive observations separated by 1 day (i.e., taken on consecutive nights) and in all three cases variability was not detected at the >10% level. The shortest time interval over which a >10% change in $W_{\text{H}\alpha}$ is detected is 3 days. Specifically, between MJD 3066 and 3069, $W_{\text{H}\alpha}$ changed by +6 Å (from $\sim$−40 Å), a change of ~15%. We conclude that, from the limited data available, variability in $W_{\text{H}\alpha}$ is present in V1647 Ori on timescales as short as 3 days.

### 3.3.3. Hα Equivalent Width in the Decline Phase

After the period of relatively constant $W_{\text{H}\alpha}$ (2004 February to 2005 September), $W_{\text{H}\alpha}$ begins to change dramatically as the major photometric decline phase starts. By the time the source has faded to close to its pre-outburst brightness (2006 February), $W_{\text{H}\alpha}$ has increased to about $\sim$−100 Å. The most obvious cause for such a dramatic change is the decrease in continuum emission as the source faded. From 2005 September to 2006 February the observed $r'$-band magnitude decreased by $\sim$4 mag or a factor 40. If the Hα flux remained constant during this period then we would expect $W_{\text{H}\alpha}$ to have increased by a factor of 40 from $\sim$−30 Å to $\sim$1200 Å. Since the increase in $W_{\text{H}\alpha}$ was measured to be only a factor $\sim$3.3, we can conclude that the Hα emission flux must have significantly decreased from 2005 September to 2006 February. We estimate this decline to be about a factor of $\sim$12. If the Hα emission is proportional to the mass accretion rate then this decline would correspond to an order of magnitude decrease in accretion from 2005 September to 2006 February. Using the relationship between near-IR H1 emission line flux and accretion rate defined by Muzerolle et al. (1998a), Acosta-Pulido et al. (2007) estimated that the mass accretion rate declined from $\sim$5 x $10^{-6}$ $M_\odot$ yr$^{-1}$ in 2004 March to $\sim$5 x $10^{-7}$ $M_\odot$ yr$^{-1}$ in 2006 May. This factor 10 reduction in accretion rate is consistent with the 12× reduction in Hα emission flux we observe between similar dates.

### 3.3.4. Full Width Zero and 2.5% Intensity

The FWZI of Hα emission can be considered a measure of twice the maximum velocity Hα emitting gas attains in the process of line creation. However, we note that this quantity is not just affected by bulk gas motion, since line broadening can occur. Muzerolle et al. (1998) suggested that Stark broadening could be an important mechanism in producing very large line widths. Muzerolle et al. (2001) presented model Hα line profiles under the influence of Stark broadening by considering a realistic range of physical parameters related to the accretion...
process, magnetosphere size, inclination angle, and stellar mass and temperature. A comparison of Hα line profiles both with and without Stark broadening (and continuum opacity) showed that this effect can be a significant contributory factor with an increase of up to 2.5× in FWZI. Their consideration of such broadening resulted in line profiles that better matched those observed in T Tauri stars. Stark broadening requires optically thick conditions and very high densities ($n_H < 10^{12}$ cm$^{-3}$), both of which are found in post-shock regions of hot accretion shocks (Hartigan et al. 1991; Valenti et al. 1993).

As mentioned above, one problem with studying FWZI is S/N. Poor S/N can mask the true wavelength at which line emission fades to the continuum level. In the case of V1647 Ori, this is particularly important in the later spectra when the source had faded below $r' = 20$. Table 5 shows the S/N at both $t_{\text{max}}$ and in the adjacent continuum for all spectra. FWZI is likely a good measure of velocity width from 2004 February to the summer of 2005 where the S/N in the continuum is always >20. However, after this time the continuum S/N drops to <5. Alternatively, FW2.5% of Hα emission has been considered (by RPL96) to be a measure of line width less sensitive to S/N since it does not require a precise definition of continuum level. This is due to FW2.5% being of a larger signal level and hence insensitive to continuum S/N issues. RPL96 observed a
### Table 5
V1647 Ori Hα Line Profiles

| MID  | S/Nb | Wc  | λpeak | Δλpeak/fλpeak | FWZI  | FW2.5%  | FW10%  | ΔV  | nchar |
|------|------|------|--------|---------------|--------|---------|---------|------|-------|
| 3049 | 252(34) | −31.4 ± 2.0 | 6563.2 ± 0.1 | +0.4(18) | 1392 ± 30 | 968 ± 30 | 534 ± 30 | 552 ± 30 | 516 ± 20 |
| 3074 | 400(39) | −50.9 ± 2.0 | 6563.0 ± 0.1 | +0.2(9) | 1384 ± 30 | 1030 ± 30 | 562 ± 30 | 270 ± 30 | 294 ± 20 |
| 3257 | 282(42) | −24.1 ± 2.0 | 6562.3 ± 0.1 | +0.5(23) | 958 ± 30 | 664 ± 30 | 460 ± 30 | 308 ± 30 | 230 ± 20 |
| 3273 | 352(10) | −18.0 ± 1.0 | 6563.3 ± 0.1 | +0.5(24) | 960 ± 10 | 592 ± 30 | 440 ± 10 | 255 ± 20 | 210 ± 10 |
| 3284 | 278(36) | −29.5 ± 2.0 | 6562.5 ± 0.1 | −0.3(14) | 1208 ± 30 | 822 ± 30 | 564 ± 30 | 273 ± 30 | 220 ± 20 |
| 3322 | 261(41) | −22.1 ± 2.0 | 6562.7 ± 0.1 | −0.1(5) | 1002 ± 30 | 660 ± 30 | 450 ± 30 | 332 ± 20 | 208 ± 20 |
| 3351 | 212(29) | −29.2 ± 2.0 | 6562.8 ± 0.1 | 0.0(0) | 1056 ± 30 | 778 ± 30 | 526 ± 30 | 248 ± 30 | 213 ± 20 |
| 3378 | 361(49) | −27.1 ± 2.0 | 6563.2 ± 0.1 | +0.4(18) | 1180 ± 40 | 856 ± 30 | 554 ± 40 | 238 ± 30 | 167 ± 20 |
| 3612 | 160(23) | −21.2 ± 2.0 | 6564.8 ± 0.1 | +0.2(9) | 840 ± 40 | 740 ± 30 | 476 ± 40 | 158 ± 30 | 150 ± 20 |
| 3638 | 108(13) | −29.2 ± 2.0 | 6562.5 ± 0.1 | −0.3(14) | 930 ± 50 | 714 ± 30 | 464 ± 50 | 112 ± 30 | 144 ± 20 |
| 3656 | 141(17) | −30.2 ± 2.0 | 6563.4 ± 0.1 | +0.6(27) | 1190 ± 50 | 864 ± 30 | 512 ± 50 | 95 ± 30 | 142 ± 20 |
| 3692 | 38(4) | −37.2 ± 2.0 | 6563.6 ± 0.1 | +0.8(37) | 970 ± 80 | 858 ± 30 | 518 ± 80 | 109 ± 60 | 127 ± 50 |
| 3701 | 29(2) | −64.2 ± 2.0 | 6563.4 ± 0.1 | +0.6(27) | 688 ± 80 | 604 ± 30 | 458 ± 80 | 98 ± 60 | 108 ± 50 |
| 3729 | 40(3) | −64.9 ± 2.0 | 6563.7 ± 0.1 | +0.9(41) | 876 ± 100 | 704 ± 30 | 536 ± 100 | <100 | − |
| 3740 | 39(3) | −66.6 ± 2.0 | 6563.9 ± 0.1 | +1.1(50) | 868 ± 100 | 680 ± 30 | 598 ± 100 | <100 | − |
| 3782 | 29(1) | −98.7 ± 2.0 | 6564.4 ± 0.1 | +1.6(73) | 770 ± 100 | 730 ± 30 | 632 ± 100 | <100 | − |
| 4152 | 15(1) | −136.2 ± 2.0 | 6565.4 ± 0.1 | +2.6(119) | 822 ± 100 | 762 ± 30 | 756 ± 100 | <100 | − |

Notes.

a Modified Julian Date. 2450000+.
b S/N on the Hα emission peak and, in parentheses, in the adjacent continuum.
c Equivalent width of emission component calculated from the continuum levels on the blue and red edges of the main Hα emission peak. The uncertainties are estimated from the spread of repeat measurements.
d The wavelength of peak Hα emission, λpeak.
e Offset of peak Hα emission from 6562.8 Å in Å and, in parentheses, km s⁻¹.
f The FWZI of the Hα emission in km s⁻¹. The velocity offset to the continuum on the red side of 6562.8 Å was measured then doubled to obtain the FWZI value. The uncertainties are estimated from the spread of repeat measurements.
g The full-width intensity of the Hα emission in km s⁻¹ measured to a signal of I_max/40 (2.5%), where I_max is the peak line intensity. The width is measured on the red side of 6562.8 Å and was doubled to obtain the FW2.5% value. The uncertainties are estimated from the spread of repeat measurements.
h The full-width intensity of the Hα emission in km s⁻¹ measured to a signal of 10% of the peak intensity, I_max. The width is measured on the red side of 6562.8 Å and was doubled to obtain the FW10% value. The uncertainties are estimated from the spread of repeat measurements.
i The width of the Hα absorption feature in km s⁻¹. The uncertainties are estimated from the spread of repeat measurements.
jj The velocity offset (in km s⁻¹) of the deepest absorption from 6562.8 Å, the rest wavelength of Hα. The uncertainties are estimated from the spread of repeat measurements.
jk Keck/HIRES spectrum. Profile not shown in Figures 20 and 21
kl ± 0.6(27) 1190 ± 1)) (km s⁻¹) (FW10% value).

significant number (63) of young low- and intermediate-mass stars at high spectral resolution (R ∼ 50,000) and presented Hα line profiles for each. The profiles were divided into four classes (Types I-IV) depending on the asymmetry of the line profile, and the location and depth of any associated absorption. The reader is referred to RPL96 for the detailed definitions of the classes.

The changes occurring in FWZI and FW2.5% with time (middle-left panel of Figure 22) mirror each other reasonable well from 2004 February to the middle of 2005. This suggests that our measurement of FWZI is most likely accurate. In the above time period, FW2.5% is ∼1.4× smaller than FWZI. Subsequently, however, the FWZI and FW2.5% values become much more similar suggesting that our values of FWZI are likely underestimates due to poor S/N and that the trend seen in FWZI is not purely intrinsic to the star. If we measure the FWZI and FW2.5% of a pure Gaussian profile we find that FW2.5% is a factor of 1.515 smaller than FWZI. Hence, the red wing of Hα in V1647 Ori between 2004 February and mid-2005 have FWZI and FW2.5% values of close to the Gaussian ratio value. We have therefore multiplied the FW2.5% values from 2005 August onwards by a factor of 1.515 and these are displayed in Figure 22 (middle-left panel) as the FWZI values.

FW2.5% decreases during 2004 from a maximum close to 1000 km s⁻¹ (FWZI ~ 1400 km s⁻¹) to ~750 km s⁻¹ (FWZI ~ 1150 km s⁻¹). After this time, the decrease in FW2.5% ceases and from 2005 onwards it remains at ~750 km s⁻¹ (scaled FWZI ~ 1150 km s⁻¹). Hence, through the major decline phase of V1647 Ori until it reaches its pre-outburst brightness, FW2.5% remains approximately constant. We note that the values encountered in V1647 Ori throughout the eruption are considerably larger (by a factor ∼3×) than those found in CTTs suggesting that accretion activity has not declined completely to a quiescent level.

The range of values of FW2.5% seen in V1647 Ori is 750–1500 km s⁻¹. Figure 9 of RPL96 shows that of their 63 sources, the peak in the distribution of (red wing) FW2.5% values lies at ~600 km s⁻¹ with only three stars showing values greater than 1000 km s⁻¹. Hence, even when V1647 Ori had faded to its pre-outburst brightness, its FW2.5% value remained larger than the RPL96 distribution peak. The three sources that RPL96 found to have FW2.5% > 1000 km s⁻¹ were R Mon, FW Cha, and AS 353A. These sources were all classified as III-B5 in the RPL96 scheme possessing blueshifted absorption. In total, 15
of their 63 sources showed type III-B profiles. The Hα profiles of V1647 Ori suggests that its RPL96 classification evolved with time from (1) IV-B-To in 2004 February/March, to (2) III-B in 2004 September/October, to (3) IV-B in 2004 November/December, to (4) III-B in 2005 January through to at least 2005 October and possibly beyond to 2007 February (although the S/N of the spectra are too poor to be definite).

3.3.5. Absorption Width and Velocity Offset

The width of the blueshifted absorption feature (herein termed $\Delta V$) and the velocity of the center of the blueshifted absorption (herein termed $v_{\text{char}}$) gives us information on the velocity field of the outflowing absorbing gas. Specifically, we consider that we can associate $\Delta V$ with the velocity dispersion in the ejected material, and $v_{\text{char}}$ with its “characteristic” velocity.

The temporal variability of $\Delta V$ (middle-right panel of Figure 22) is considerably different from that of the Hα emission component. A very rapid decline in absorption width is seen between our first two observations taken in 2004 February ($\Delta V \sim 600$ km s$^{-1}$) and 2004 March ($\Delta V \sim 300$ km s$^{-1}$). After this time, $\Delta V$ is approximately constant to the end of our winter 2004–2005 monitoring season. At the start of the fall 2005 observing period, $\Delta V$ has decreased to $\sim 100$ km s$^{-1}$ and there is some suggestion that this decrease was related to the main photometric decline. During the fall/winter months of 2005, $\Delta V$ remains approximately constant although the S/N of the spectra is poor. The final observation point, February 2007, also had very low S/N but the indication is that the value of $\Delta V$ had changed little.

The velocity offset of the deepest absorption, $v_{\text{char}}$ (bottom-left panel, dashed line), mirrors $\Delta V$ showing a similar large change between the first two observations (from $\sim 500$ km s$^{-1}$ to $\sim 300$ km s$^{-1}$). This is followed by a slower decline (to $\sim 200$ km s$^{-1}$) between the spring and fall of 2004. In the fall of 2005 $v_{\text{char}}$ had again decreased, this time by a factor of $\sim 2$ (to $\leq 100$ km s$^{-1}$). By early 2006, when V1647 Ori had returned to its pre-outburst brightness, the S/N of the spectra and the weakness of the absorption made a determination of $v_{\text{char}}$ difficult and all we can say is that it was $\leq 100$ km s$^{-1}$.

3.3.6. Peak Emission Wavelength

The wavelength of peak Hα emission is a rather complex quantity since it is dependent not only on the motion of the emitting material, but also on the characteristics of any associated absorption. For example, if a blueshifted wind absorbs the blue wing of the emission line then the wavelength of peak emission can shift to the red. Alternatively, if the blueshifted absorption is weak and we measure a wavelength shift in peak emission, then we could interpret this as evidence for bulk motion of the emitting gas. Multiple absorption components, as we saw in Figure 18, also complicates the interpretation. In light of the above discussion, we refrain from any ad hoc interpretation of the changes in $\lambda_{\text{peak}}$ and merely relate the variations observed.

The final panel of Figure 22 shows the wavelength of peak Hα emission ($\lambda_{\text{peak}}$, bottom-right panel) over the monitoring period. Here, we see a peak-to-peak variability of $\sim 3$ Å ($\sim 140$ km s$^{-1}$) from 2004 February to 2007 February. Soon after outburst (2004 February/March), the peak in emission was measured to be slightly to the red of our assumed Hα rest wavelength (6562.8 Å). By the fall of 2004, $\lambda_{\text{peak}}$ had shifted to the blue by about 1 Å ($\sim 45$ km s$^{-1}$). During the fall/winter of 2004, the peak had shifted back to its original wavelength (of 2004 February/March). When our monitoring commenced again in the fall of 2005, $\lambda_{\text{peak}}$ was at about the same wavelength as in spring 2005 but soon after it shifted briefly to the blue and then rapidly to the red and by the spring of 2006 it was located redward of the rest wavelength by $\sim 1.7$ Å ($\sim 80$ km s$^{-1}$). This trend continued until our last observation, in 2007 February, when $\lambda_{\text{peak}}$ was at $+2.5$ Å ($\sim 120$ km s$^{-1}$).

3.4. Hα Profile Modeling

It is generally agreed that the presence of complex Hα emission structure with blueshifted absorption in the spectrum of a young star is the result of the combination of emission from magnetospheric accretion and absorption from strong, dense stellar winds. The recent papers by Muzerolle et al. (1998), Alencar et al. (2005), Kurosawa et al. (2006), and Bouvier et al. (2007) discuss models involving such processes and their fit to observations of low-mass stars. It is clearly important to consider such numerical simulations to available data sets to allow the derivation of physical parameters such as outflow and accretion rates, and test the validity of the specific models. However, due to the complexity of such analyses in regions of high accretion (R. Kurosawa 2007, private communication), what follows will concentrate on a qualitative understanding of the changes observed.

4. PHYSICAL COMPONENTS OF THE OUTBURST

The eruptive event that commenced between 2003 October 23 and 2003 November 15 (Briceno et al. 2004) resulted in the dramatic brightening of V1647 Ori and the appearance of what is now known as McNeil’s Nebula. The event was apparently extremely violent since about three months after the outburst occurred (when our first observations were taken) the star had optically brightened by a factor of $\sim 100$ and produced a P Cygni Hα profile exhibiting both a very high velocity ($\sim 600$ km s$^{-1}$), blueshifted, absorption trough (indicative of the presence of a dense wind) and strong emission with an enormous velocity width (i.e., FWZI $\sim 1500$ km s$^{-1}$—indicative of magnetospheric accretion and perhaps Stark broadening). How these Hα spectral features evolved over the first three months of the eruption is unknown, however, they are likely to be at least similar in nature to those observed at three months old if not even more extreme. Below, we summarize what we can deduce from our results regarding the properties and evolution of the wind and accretion phenomena.

4.1. The Fast Wind

At the three month mark of the eruption, a strong wind existed producing two minima in the blueshifted absorption (see Figure 21). Such a double-trough absorption profile was evident at times, for example, in the spectra of V1057 Cyg, a classical FUor (see Bastian & Mundt 1985). However, the two components of the wind disappeared between our first and second observations on UT 2004 February 14 and UT 2004 March 10, respectively. The optical spectrum of Ojha et al. (2005) from UT 2004 February 22 did not show such features either. The two minima in the Hα absorption on UT 2004 February 14, were at velocities of $v = -270$ and $-530$ km s$^{-1}$ with respect to the rest wavelength of Hα (6562.8 μm). On UT 2004 March 10, the single absorption minimum was at $v = -350$ km s$^{-1}$. A double-peaked absorption could be

---

6 P Cygni profile with absorption at sufficiently high velocity absorption to absorb continuum flux beyond the wing of the Hα emission.
associated with multiple shell-like wind components ejected with different velocities or with the geometry and motion of the emission region. A continuous slowing of the wind is indicated by the $v_{\text{char}}$ values plotted in Figure 22 where the velocity of the absorption minimum declined from a maximum of $-530 \text{ km s}^{-1}$ to a minimum of $-150 \text{ km s}^{-1}$. The width of the absorption component, $\Delta V$, also declined over this period (from 560 to 100 km s$^{-1}$) indicating a reduction in velocity dispersion in the wind and suggesting that the wind decelerated as it expanded away from V1647 Ori. Such behavior is perhaps consistent with the stochastic wind models of Grinin & Mitskevich (1992) and Mitskevich et al. (1993). In these models, the wind is nonisotropic and clumpy and decelerates at large distances from the young star.

4.2. The Emission Region

During the period of rapid decline of $v_{\text{char}}$, i.e., between 2004 February 14 and March 10, the FWZI of the H$\alpha$ emission remained relatively constant at $\sim 1500 \text{ km s}^{-1}$. As we saw above, such a value is somewhat larger than typically found in accreting classical T Tauri stars (from the work of RPL96). Additional examples of this are RW Aur, which exhibited a FWZI of $\sim 1000 \text{ km s}^{-1}$ in the spectra of Alencar et al. (2005), and some sources studied by Muzerolle et al. (1998) and Alencar & Basri (2000) where a typical FWZI was found to be in the range of 600–1000 km s$^{-1}$. The FWZI did not show significant decline until sometime between our observation in 2004 March and our next observation in 2004 September. Over this period the FWZI had reduced by around 500 km s$^{-1}$ (to $\sim 1000 \text{ km s}^{-1}$). During the period of the major photometric decline, from 2005 August to 2006 February, the FWZI remained in the range 1000–1400 km s$^{-1}$, i.e., the dramatic change in optical brightness of V1647 Ori did not appear to be strongly correlated with the FWZI of the H$\alpha$ emission. This argues that the optical fading and the H$\alpha$ emission are not the result of the same physical processes.

As we related above, contributions to the FWZI of the H$\alpha$ emission probably arise in both bulk gas motion and Stark broadening. The region from which the H$\alpha$ arises therefore must be extremely dense suggesting that it is the post-shock region of the accretion flow. The change in optical colors as the major photometric decline phase began (they became redder) suggests that there was a reduction in blue continuum from the accretion shock region. Briceño et al. (2004) observed the spectrum of V1647 Ori using reflected light from McNeil’s Nebula and found it to correspond to that from an early B-type source, i.e., the hot accretion flow. Correspondingly, the reduction in blue continuum likely resulted from a decline in post-shock temperature and density, probably due to a slowing of the accretion rate. The decline in accretion rate was estimated, using near-IR H lines, to be an order of magnitude between 2004 March and 2006 May (Acosta-Pulido et al. 2007).

5. OUTBURST TIMELINE

We now attempt to piece together the timeline of events over the outburst period and relate the photometric and spectroscopic variations discussed above.

1. Prior to the outburst, V1647 Ori appeared to be in an unstable state with variations in optical brightness of up to $\sim 2$ mag. It appeared that the main outburst event was initiated from an photometrically elevated state existing for at least five years prior to the main eruption, i.e., from the first Briceño et al. (2004) and McGehee et al. (2004; SDSS) observations.

2. The outburst began between 2003 late-July and late-October. Unfortunately, there are only a handful of photometric measurements for the first three+ months of the event. These were observed serendipitously during an Orion OB1 photometric monitoring program (Briceño et al. 2001). The rise time to peak brightness was short, 120 days, with a change in optical (I-band) brightness of 3.6 mag.\footnote{V1647 Ori had already brightened by $\sim 1$ mag prior to this but it is unknown when this occurred due to lack of observations during this period (Briceño et al. 2004).}

3. At the time of our first observations, the first taken on the source after its discovery by McNeil (2004), the star had brightened from its pre-outburst level by over 5 mag in the optical and over 2 mag at 2 $\mu$m. It had optical spectral features indicative of a massive accretion burst producing a fast, dense stellar wind creating high-speed blueshifted P Cygni-type absorption on strong H$\alpha$ emission. In the NIR, V1647 Ori possessed strong emission in Br$\gamma$, Na i, and the $\nu = 2–0 \text{ CO overtone bandheads, all indications were that the emission region was hot, extremely dense, and in an excited kinematic state. The optical brightness is attributed to a combination of the star being revealed after dust in the immediate circumstellar environment of the object was sublimated (Aspin et al. 2009b) together with a luminosity increase due to the addition of significant hot accretion flux. The brightening in the NIR is perhaps more likely the effect of irradiation and subsequent heating of the inner regions of the circumstellar disk, beyond the sublimation radius, by UV flux from the accretion flow. However, it is clear that the heating did not propagate through to outer regions of the disk since the submm mm$^{-1}$ flux remained unchanged.

4. Over the next 10 months, V1647 Ori remained within half a magnitude of its outburst optical/NIR brightness although there was seemingly stochastic variability at the level of around 1 mag peak-to-peak. Statistically significant brightness fluctuations were seen on periods as short as a few days. During this so-called “high-plateau” phase, the optical spectrum of V1647 Ori showed strong P Cygni H$\alpha$ emission on a relatively stable continuum. This suggests that the accretion process continued unabated over this period with only minor variations occurring. This is supported by the reasonably constant FWZI of H$\alpha$ ($\sim 1500 \text{ km s}^{-1}$), indicating that the emission was taking place in very dense gas, and the derived accretion rate from NIR line emission. During this same period, the wind quickly slowed (from $v_{\text{char}} \sim 600–200 \text{ km s}^{-1}$). Also, the velocity dispersion in the wind mirrored $v_{\text{char}}$ in that it quickly declined (from $\Delta V \sim 600–200 \text{ km s}^{-1}$). During the rapid decline in wind velocity and velocity dispersion, weak shock-excited emission lines of [S ii] appeared in the optical spectrum. These remained approximately constant to the end of the monitoring period. During the summer months of 2004, an [O i] emission line appeared and strengthened for about a year. This may be the result of the formation of a partly obscured bipolar-like outflow and its formation approximately corresponded to the period of the high-plateau. Subsequently, the [O i] line declined in intensity which corresponded to the rapid decline in brightness of V1647 Ori.

5. At the start of the rapid decline phase of the optical flux (in the fall of 2005, some six months after our last plateau-phase observation), V1647 Ori had already shown...
a significant reduction in Hα FWZI (~1500 km s\(^{-1}\) to ~800 km s\(^{-1}\)), wind velocity (\(v_{\text{rad}}\)~300 km s\(^{-1}\) to ~200 km s\(^{-1}\)), wind velocity dispersion (\(\Delta V \sim 250\) km s\(^{-1}\) to ~150 km s\(^{-1}\)), and accretion rate (a factor of 10). This suggests that the density of the region of Hα emission declined and the wind slowed. Two of the aforementioned parameters (FWZI and \(\Delta V\)) remained approximately constant to the end of our monitoring period (in 2006 February) when the source had faded by ~4 mag and returned to its pre-outburst optical brightness. The only significant change was in wind velocity which continued to decline (\(\delta v \sim 250\) km s\(^{-1}\) to ~100 km s\(^{-1}\)) into early 2006 implying that the wind was still decelerating.

6. In the period after 2006 February, beyond the time period investigated here, we obtained one further set of observations that proved rather interesting. These were published in Aspin et al. (2008) and were obtained approximately one year after the last observation presented above. In 2007 February, the optical brightness of V1647 Ori remained close to its 2006 February value with \(r' = 23.3\). Little evidence was seen for blueshifted Hα absorption suggesting that the wind had continued to weaken over the intervening year. The FWZI of Hα remained at ~800 km s\(^{-1}\) implying that the accretion process had perhaps settled into a quasi-steady state. The Hα line profile calculations by Muzerolle et al. (2001) discussed above, gave values of FWZI of this order (see their Figures 8, 11, and 12) and produced a good match to observations of some CTTSs (e.g., BP Tau with FWZI ~800 km s\(^{-1}\)). In addition, the February 2007 data showed that shock-excited [S ii] emission was still present, as was weak [O i] emission with a weak blueshifted component.

We are grateful to G. H. Herbig for providing the HIRES spectrum of V1647 Ori presented above. We also thank C. Briceño for access to his 2004 February 18 spectrum of V1647 Ori. We additionally thank D. Finkbeiner and P. McGeehee for assistance with accessing SDSS data, and the latter for providing numerical model information. This work is based on observations obtained at the Gemini Observatory (under program identifications GN-2004A-DD-3, GN-2004B-Q-28, GN-2005B-Q-1), which is operated by the Association of Universities for Research in Astronomy, Inc., under a cooperative agreement with the NSF on behalf of the Gemini partnership: the National Science Foundation of Argentina, the National Aeronautics and Space Administration, the Japanese Monbukagakusho, the Max Planck Society, and the National Aeronautics and Space Administration. The SDSS Web site is http://www.sdss.org/. The SDSS is managed by the Astrophysical Research Consortium for the participating institutions. The participating institutions are the American Museum of Natural History, Astrophysical Institute Potsdam, University of Basel, Cambridge University, Case Western Reserve University, University of Chicago, Drexel University, Fermilab, the Institute for Advanced Study, the Japan Participation Group, Johns Hopkins University, the Joint Institute for Nuclear Astrophysics, the Kavli Institute for Particle Astrophysics and Cosmology, the Korean Scientist Group, the Chinese Academy of Sciences (LAMOST), Los Alamos National Laboratory, the Max-Planck-Institute for Astronomy (MPIA), the Max-Planck-Institute for Astrophysics (MPA), New Mexico State University, Ohio State University, University of Pittsburgh, University of Portsmouth, Princeton University, the United States Naval Observatory, and the University of Washington. This research has made use of NASA’s Astrophysics Data System Bibliographic Services. This research has made use of the SIMBAD database, operated at CDS, Strasbourg, France. We are most fortunate to have the opportunity to conduct astronomical observations from the sacred mountain Mauna Kea. During this research, C.A. was supported in part by NASA through the American Astronomical Society’s Small Research Grant Program. B.R. acknowledges partial support from the NASA Astrobiology Institute under Cooperative Agreement No. NNA04CC08A.

**REFERENCES**

Abrahám, P., Kóspal, A., Csizmadia, S., Moór, A., Kun, M., & Stringfellow, G. 2004a, A&A, 419, L39
Abrahám, P., et al. 2004b, A&A, 428, 89
Abrahám, P., et al. 2006, A&A, 445, 13
Acosta-Pulido, J. A., et al. 2007, AJ, 133, 2020
Aleren, S. H. P., & Basri, G. 2000, AJ, 119, 1881
Aleren, S. H. P., Basri, G., Hartmann, L., & Calvet, N. 2005, A&A, 440, 595
Andrews, S. M., & Williams, J. P. 2004, ApJ, 610, L45
Aspin, C., Barbieri, C., Boschi, F., Di Mille, F., Rampazzi, F., Reipurth, B., & Tsu...
