Periodic Variations in the Residual Eclipse Flux and Eclipse Timings of Asynchronous Polar V1432 Aql: Evidence of a Shifting Threading Region

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ABSTRACT

We report the results of a twenty-five-month photometric campaign studying V1432 Aql, the only known eclipsing, asynchronous polar. Our data show that both the residual eclipse flux and eclipse O–C timings vary strongly as a function of the spin-orbit beat period. Relying upon a new model of the system, we show that cyclical changes in the location of the threading region along the ballistic trajectory of the accretion stream could produce both effects. This model predicts that the threading radius is variable, in contrast to previous studies which have assumed a constant threading radius. Additionally, we identify a very strong photometric maximum which is only visible for half of the beat cycle. The exact cause of this maximum is unclear, but we consider the possibility that it is the optical counterpart of the third accreting polecap proposed by Rana et al. (2005). Finally, the rate of change of the white dwarf’s spin period is consistent with it being proportional to the difference between the spin and orbital periods, implying that the spin period is approaching the orbital period asymptotically.

Subject headings: accretion, accretion disks — binaries: eclipsing — novae, cataclysmic variables — stars: individual (V1432 Aql, RX J1940.1-1025) — stars: magnetic field — white dwarfs

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1. Introduction

Cataclysmic variables (CVs) are interacting binary systems in which a low-mass star—usually a red dwarf—overfills its Roche lobe and transfers mass onto a white dwarf (WD). Hellier (2001) and Warner (1995) offer excellent overviews of these intriguing systems. In a subset of CVs known as polars, the exceptionally strong magnetic field of the WD synchronizes the WD’s spin period with the orbital period of the binary (see Cropper (1990) for a comprehensive review). The accretion stream from the secondary star follows a ballistic trajectory toward the WD until the magnetic pressure matches the stream’s ram pressure. When this occurs, a threading region forms in which the accretion stream couples onto the WD’s magnetic field lines, and the captured material is then channeled onto one or more accretion regions near the WD’s magnetic poles. The impact of the stream creates a shock in which the plasma is heated to X-ray-emitting temperatures, so polars are far brighter in X-ray wavelengths than ordinary non-magnetic CVs. In addition to X-rays, the accretion region produces cyclotron emission in the optical and in the infrared, the detection of which is a defining characteristic of polars.

Eclipses of the WD have provided great insight into polars. Because a polar has no accretion disk, an eclipsing polar will generally exhibit a two-step eclipse: a very sharp eclipse of the compact (∼white dwarf radius) cyclotron-emitting region, followed by a much more gradual eclipse of the extended accretion stream (see, e.g., Harrop-Allin et al. (1999) for an eclipse-mapping study of HU Aqr). When the accretion rate is high, the WD photosphere makes only a modest contribution to the overall optical flux, overshadowed by the two accretion-powered components mentioned above. Eclipsing polars also make it possible to determine the orientation of the magnetic axis with respect to the secondary. In HU Aqr, the orientation of the dominant magnetic pole leads the line of centers of the binary by about 45° (Harrop-Allin et al. 1999) while in DP Leo, another eclipsing polar, the equilibrium orientation leads the line of centers by 7° ± 3° but with a long-term oscillation with an amplitude of ∼25° (Beuermann et al. 2014).

In at least four polars, the WD has a spin period which differs by no more than several percent from the orbital period. As Stockman, Schmidt, & Lamb (1988) describe, these asynchronous polars probably result from novae, which break the system’s synchronism by causing the primary to lose mass and to interact with the secondary. Because the WD’s magnetic field will gradually resynchronize the spin period with the orbital period,

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1In addition to the subject of this study (V1432 Aql), three other polars are incontrovertibly asynchronous: BY Cam, V1500 Cyg, and CD Ind. There are at least two additional candidates at the time of writing: V4633 Sgr (Lipkin & Leibowitz 2008) and CP Pup (Bianchini et al. 2012).
asynchronous polars are short-lived phenomena. For example, Schmidt & Stockman (1991) detected a derivative in the WD spin period in V1500 Cyg (which had undergone a nova in 1975) and estimated that the system would approach resynchronization about 150 years after the publication of their study.

Interestingly, in each of these four confirmed asynchronous polars, the threading process is inefficient compared to fully synchronous systems. For example, a Doppler tomogram of V1432 Aql showed an extended accretion curtain (Schwope 2001), a finding which is possible only if the accretion stream can travel significantly around the WD. X-ray observations of V1432 Aql also indicate that the accretion stream travels most of the way around the WD before it is fully threaded onto the magnetic field lines (Mukai et al. 2003). As for the other confirmed asynchronous polars, there is mounting evidence that the accretion flow can significantly extend around the WD. In CD Ind, the accretion stream appears to thread onto the same magnetic field line throughout the beat cycle, requiring that the stream be able to travel around the WD (Ramsay et al. 1999). With regard to V1500 Cyg, Schmidt & Stockman (1991) argued that the smooth sinusoidal variation of the polarization curve was consistent with the infalling stream forming a thin accretion ring around the WD. More recently, Litvinchova, Pavlenko, & Shugarov (2010) detected evidence that this accretion ring is fragmented, periodically reducing the irradiation of the donor star by the hot WD. In the remaining system, BY Cam, Doppler tomograms show that the accretion curtain extends over \(\sim 180^\circ\) in azimuth around the WD, requiring a similar extent of the accretion stream (Schwarz et al. 2005). By contrast, the accretion stream in synchronous systems is captured before it can travel around the WD (e.g. Schwepe, Mantel, & Horne 1997). Although a sample size of four is small, it is remarkable that in each of the confirmed asynchronous polars, the threading process is so inefficient that the accretion stream can travel much of the way around the WD.

1.1. V1432 Aql

V1432 Aql (= RX J1940.1-1025) is the only known eclipsing, asynchronous polar and was identified as such by Patterson et al. (1995) and Friedrich et al. (1996). There are two stable periodicities in optical photometry of V1432 Aql. The first (12116 seconds) is the orbital period, which is easily measured from the timings of the eclipses of the WD by the secondary. Initially, the nature of the eclipses was unclear; Patterson et al. (1995) argued that the secondary was the occulting body, but Watson et al. (1995) contended that a dense portion of the accretion stream was the culprit. Much of the confusion was attributable to the presence of residual emission lines and X-rays throughout the eclipses, as well as
the variable eclipse depth. Since X-rays in polars originate on or just above the WD’s surface, the presence of flux at these wavelengths throughout the eclipse was inconsistent with occultations by the donor star. Additionally, there was considerable scatter in the eclipse timings, and the system’s eclipse light curves did not show the rapid ingresses and egresses characteristic of synchronous polars ([Watson et al.] 1995). However, [Mukai et al.] (2003) resolved the dispute with high-quality X-ray observations which revealed that the residual X-ray flux was actually contamination from a nearby Seyfert galaxy, thereby ruling out the stream as the source of the eclipses.

The second periodicity (≈ 12150 seconds) is the spin modulation of the WD. In optical photometry, this periodicity manifests itself in several ways. In particular, at $\phi_{\text{spin}} = 0$, the WD is occulted by material accreting onto one of the magnetic poles, producing a broad “spin minimum” ([Friedrich et al.] 1996). Analyses of the spin minima have revealed several fascinating insights into V1432 Aql. For example, [Geckeler & Stauber] (1997) undertook an O–C study of the timing residuals of the spin minima and managed to detect a decrease in the WD spin period, indicating that the system is resynchronizing itself. They also measured a cyclical variation in the timings of the spin minima, caused by (1) a longitudinal offset between the magnetic pole and its corresponding accretion region on the WD’s surface and (2) the accretion stream threading onto different magnetic field lines throughout the spin-orbit beat period ($P_{\text{beat}} = |P_{\text{orb}}^{-1} - P_{\text{spin}}^{-1}|$). Using these timings and a dipole accretion model, the authors managed to constrain the combined effect of the threading radius and the colatitude of the magnetic axis on the WD, but they could not constrain these parameters individually. [Staubert et al.] (2003) applied the methodology of [Geckeler & Stauber] (1997) to a larger dataset and refined the results of the earlier paper.

A critical concept which emerges from the literature is the beat period between the spin and orbital periods. The beat period is simply the amount of time that it takes for the WD (and its magnetic field) to rotate once as seen from the perspective of the donor star. As [Geckeler & Stauber] (1997) first demonstrated, the accretion stream will interact with different magnetic field lines as the system progresses through its beat period, a foundational principle which informs our analysis throughout this paper.

V1432 Aql is especially suitable for long-term study because its long-term brightness has remained constant not only in our own observations but in data from the American Association of Variable Star Observers as well. While many polars alternate unpredictably between bright and faint states due to changes in the mass-transfer rate, V1432 Aql does not appear to do so.

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2 www.aavso.org
We supplement these previous studies by reporting the detection of stable periodicities in both the residual eclipse flux and the O–C timing residuals of the eclipses. These phenomena occur at the beat period, and we use a model to show that our observations are consistent with a threading radius whose position and altitude above the WD vary throughout the beat cycle.

In response to this study’s observational findings, one of us (DB) followed up by analyzing a different set of observations obtained by the Center for Backyard Astrophysics\(^3\) over a much longer timespan. His group’s analysis provides confirmation of the residual-flux and timing variations described in this paper while also reporting additional beat-cycle-related phenomena (Boyd et al. 2014).

2. Observations

As part of a twenty-five-month effort to study V1432 Aql’s behavior at different beat phases, six of us (CL, RM, RC, KCM, TC, and DS) obtained unfiltered, time-resolved photometry using the University of Notre Dame’s 28-cm Schmidt-Cassegrain telescope and SBIG ST-8XME CCD camera between July 2012 and July 2014. The exposure time was 30 seconds for each individual image, with an overhead time of 8 seconds per image. A total of 70 light curves, consisting of over 17,000 individual measurements, were obtained with this instrument. These observations constitute the bulk of our dataset, and their uniformity avoids the introduction of errors caused by combining unfiltered observations from different telescope-CCD combinations. CL also used the 80-cm Sarah L. Krizmanich Telescope, also located at the University of Notre Dame, to collect four time series with an SBIG STL-1001E CCD camera in July 2014. The data obtained with the Krizmanich Telescope have much higher time resolution (7-second exposures, each with a 3-second readout time, for a total cadence of 10 seconds), facilitating the study of the rapid variability during the eclipses.

In addition, MC used a 40-cm Schmidt-Cassegrain and QSI-516 CCD camera with a Johnson V filter, JU used a 23-cm Schmidt-Cassegrain and QSI-583ws CCD camera, and DB used a 25-cm Newtonian with an unfiltered SXV-H9 CCD camera. Each of them used 60-second exposures.

The timestamp for each observation was corrected to the BJD (TDB) standard (Eastman, Siverd, & Gaudi 2010). To infer V magnitudes from our unfiltered data, we used field photometry from the AAVSO to select relatively blue reference stars whose V magnitudes were used when cal-

\(^3\)http://cbastro.org/
calculating V1432 Aql’s magnitude. All magnitudes discussed in this paper are inferred V magnitudes.

One of the most obvious phenomena in the photometry is the highly variable magnitude of the system at mid-eclipse, which ranges from V \sim 16.0 to V \sim 17.5. Different eclipses also displayed strikingly different morphologies, and in Figure 1 we plot two eclipses which are representative of this variation. Such behavior is plainly at odds with normal eclipsing polars, which almost invariably have very abrupt ingresses and egresses since most of the flux originates in a small—and thus rapidly eclipsed—area on the WD. V1432 Aql’s gradual ingresses and egresses indicate that its flux originates in an extended region, and in this regard, its eclipses bear a superficial resemblance to those of CVs with accretion disks.

We measured both the time of minimum residual eclipse flux and the magnitude at mid-eclipse by fitting a fifth-order polynomial to each eclipse (see Table 1). Figure 1 demonstrates the adequacy of the fit by plotting two eclipse light curves, each fitted with a fifth-order polynomial. Since the system’s eclipses are frequently asymmetric, the time of minimum flux is not necessarily the midpoint between ingress and egress. Indeed, several eclipses were W-shaped, with two distinct minima. For these eclipses, we report the time of the deepest minimum. One particularly remarkable eclipse, observed on JD 2456843 and discussed in Section 3.3.4, had two minima of equal depth, so we report both times.

Additionally, we detected a number of spin minima. Since previous studies of the spin minima (e.g., Geckeler & Staubert (1997)) have measured minima timings by searching for the vertical axis of symmetry of each spin minimum, we fit a second-order polynomial to each spin minimum. While a higher-order polynomial would do a better job of modeling the often-asymmetric spin minima, using the second-order polynomial increases the compatibility of
our timings with those presented in other works. We report in Table 2 the timings of all clearly-detected spin minima. We did not measure spin minima which were ill-defined or which had multiple minima of comparable depth because in those instances, it was not possible to objectively identify the middle of the spin minima.

3. Analysis

3.1. Updated Orbital and Spin Ephemerides

We used $\chi^2$ minimization to determine the best-fit ephemerides for the spin and orbital periods using our data in conjunction with the published optical eclipse and spin-minima timings in Geckeler & Staubert (1997), Staubert et al. (2003), Mukai et al. (2003), and Bonnardeau (2012). Some of the timings from these studies lacked uncertainties; for those observations, we adopted the average uncertainty of all measurements which did have error estimates. Furthermore, Andronov, Baklanov, & Burwitz (2006) have made their photometry of V1432 Aql available electronically, and while their time resolution was too low for inclusion in our eclipse analysis, it was adequate for measuring the spin minima. In the interest of uniformity of analysis, we measured the spin minima in the Andronov, Baklanov, & Burwitz (2006) dataset ourselves instead of using their published timings.

The best-fit linear eclipse ephemeris is

$$T_{\text{ecl}}[\text{HJD}] = T_{0,\text{ecl}} + P_{\text{orb}}E_{\text{ecl}},$$

with $T_{0,\text{ecl}} = 2454289.51352 \pm 0.00004$ d and $P_{\text{orb}} = 0.1402347644 \pm 0.0000000018$ d cycle$^{-1}$. Even though our timestamps use the BJD standard, we report our epochs using the slightly less accurate HJD standard because the previously published data use HJD. We found no evidence of a period derivative in the orbital ephemeris, but both Bonnardeau (2012) and Boyd et al. (2014) have reported quadratic orbital ephemerides. The latter paper had a larger dataset than the one used in this study, so our non-detection of an orbital period derivative is hardly conclusive.

The spin ephemeris of Bonnardeau (2012) fits our data very well, and we offer only a modestly refined cubic spin ephemeris of

$$T_{\text{min}}[\text{HJD}] = T_{0,\text{sp}} + P_0E_{\text{spin}} + \frac{\dot{P}}{2}E_{\text{spin}}^2 + \frac{\ddot{P}}{6}E_{\text{spin}}^3,$$

where $T_{\text{min}}$ is the midpoint of the spin minimum, $T_{0,\text{sp}} = 2449638.3278(\pm0.001)$ d, $P_0 = 0.14062844(\pm0.000000023)$ d cycle$^{-1}$, $\dot{P}/2 = -8.16(\pm0.11) \times 10^{-10}$ d cycle$^{-2}$, and $\ddot{P}/6 =$
\( -9.3(\pm 1.5) \times 10^{-16} \text{ d cycle}^{-3} \). The uncertainties on these parameters were determined via the Monte Carlo method. We do not have enough observations to meaningfully search for higher-order period derivatives like those reported by Boyd et al. (2014).

While a polynomial fit accurately models the existing data, \( P_{sp} \) will likely approach \( P_{orb} \) asymptotically (P. Garnavich, private communication). If this is correct, then \( \dot{P} \) is likely proportional to the difference between \( P_{sp} \) and \( P_{orb} \) so that

\[
\dot{P} \equiv \frac{dP_{sp}}{dE_{sp}} = k(P_{sp} - P_{orb}). \tag{3}
\]

Integrating the solution to this differential equation yields an ephemeris of

\[
T_{min} = \frac{P_{sp,0} - P_{orb}}{k}(e^{kE_{sp}} - 1) + P_{orb}E_{sp} + T_0, \tag{4}
\]

where \( P_{orb} \) is the measured value and the three free parameters are \( k = -4.207(\pm 0.017) \times 10^{-6} \text{ cycles}^{-1} \), \( P_{sp,0} = 0.14062864(\pm 0.00000017) \text{ d cycle}^{-1} \), and \( T_0 = 2449638.3277(\pm 0.0010) \text{ d} \).

\( \chi^2_{red} \) is deceptively high for both the cubic ephemeris (\( \chi^2_{red} = 3.42 \)) and the exponential ephemeris (\( \chi^2_{red} = 3.40 \)) because these ephemerides assume that the spin minimum corresponds with the transit of a fixed point on the WD’s surface. Therefore, they do not take into account the cyclical shifts in the location of the accretion spot first reported by Geckeler & Staubert (1997), a point underscored by Figure 2, which plots the O–C timing residuals of the spin minima based on our cubic ephemeris. Unless a spin ephemeris attempts to model these variations and their \( \sim 1000\text{-second peak-to-peak amplitude} \), it would be difficult to achieve a more reasonable \( \chi^2_{red} \). We did not try to incorporate this periodicity into our ephemerides because it is not an actual variation in the spin period. With this caveat in mind, the comparable values of \( \chi^2_{red} \) for each ephemeris lead us to conclude that they model the data equally well, with the exponential ephemeris having a marginally better fit. Though we use the cubic ephemeris for the sake of simplicity when calculating the beat phase in Sec. 3.2, the exponential ephemeris is at least grounded in a physical theory of the resynchronization process. Moreover, in principle, the only parameter which should need to be refined in the future is the constant \( k \). By contrast, as time passes, a polynomial ephemeris might eventually require an ungainly number of terms in order to attain a satisfactory fit.

As defined by Schmidt & Stockman (1991), a first-order approximation of an asynchronous polar’s synchronization timescale is given by

\[
\tau_s = \frac{P_{orb} - P_{spin}}{\dot{P}}. \tag{5}
\]

If one assumes rather unrealistically that \( \dot{P} \) will remain constant until resynchronization, this formula provides a very rough estimate of when resynchronization will occur. If Equation 3 is
Fig. 2.— $O-C$ timing residuals for the spin minima as a function of $\phi_{beat}$. The black dataset represents the new timings which we report in Table 2, while the gray datapoints are from previously published studies as described in the text. The data are repeated for clarity. Our lack of timings from $0.0 < \phi_{beat} < 0.5$ is a consequence of the weakness of the spin minima during this half of the beat cycle.

substituted for $\dot{P}$ in Equation 5, this equation simplifies to $\tau_s = -k^{-1}$. Since $k$ is essentially a decay rate, this formula yields the amount of time necessary for the initial value (in this context, the asynchronism at $T_0$, given by $P_{spin,0} - P_{orb}$) to be reduced by a factor of $e^{-1}$. As a result, our value of $\tau_s = 237700$ spin cycles means that in the year $\sim2086$, the spin period should be $\sim12128.8$ seconds, fully 12.5 seconds longer than $P_{orb}$. Clearly, this estimate of $\tau_s$ is not an estimate of when resynchronization will actually occur. Our value of $\tau_s = 71.5 \pm 0.4$ years (with respect to August 2014) is slightly less than the values in Geckeler & Staubert (1997) and Andronov, Baklanov, & Burwitz (2006) and considerably less than Staubert et al. (2003).

It is unclear how long an exponential ephemeris will remain valid, but if ours were to hold true indefinitely, it predicts that $P_{spin}$ will approach $P_{orb}$ to within one second in the year $\sim2320$ and to within 0.1 seconds in $\sim2750$. These are not synchronization timescales as defined by Schmidt & Stockman (1991), but in the case of an exponential ephemeris, they provide a more realistic manner of extrapolating when the system might approach resynchronization. The inferred $\sim300$-year timespan necessary just to attain $P_{spin} - P_{orb} < 1$ seconds is longer than the $\sim100$-year timescales in Geckeler & Staubert (1997) and Andronov, Baklanov, & Burwitz (2006), but it is within the error bounds of the $\sim200$-year synchronization timescale announced in Staubert et al. (2003).
3.2. Determining the Beat Phase

The spin-orbit beat cycle is the key to making sense of V1432 Aql’s behavior. To calculate the beat phase ($\phi_{\text{beat}}$) of an observation is to determine the relative orientation of the WD’s magnetic field at that time. However, since the WD spin period is variable, the beat period ($P_{\text{beat}}$) changes, too. For example, Patterson et al. (1995) measured $P_{sp} = 12150$ seconds, leading to a $P_{\text{beat}} \sim 50$ days. But by 2013, the spin period had decreased, leading to a beat period of $\sim 62$ days. We pause briefly to outline the procedure that we employed to calculate $\phi_{\text{beat}}$ given the time of observation ($T$).

Since $P_{\text{beat}}$ is given by

$$P_{\text{beat}}^{-1} = |P_{\text{orb}}^{-1} - P_{\text{spin}}^{-1}|,$$

one solution is to determine the average length of the spin period ($\bar{P}_{\text{spin}}$) between $T$ and $T_0$. The first step in determining $\bar{P}_{\text{spin}}$ is to differentiate the cubic spin ephemeris from Section 3.1 with respect to the spin epoch $E$. This yields a formula for the instantaneous spin period, the average value of which is

$$\bar{P}_{\text{spin}} = \frac{1}{E_T} \int_0^{E_T} (P_0 + \dot{P}E + \frac{1}{2} \ddot{P}E^2) dE.$$ \hspace{1cm} \text{(7)}

The spin epoch $E_T$, in turn, is found by applying the cubic formula to the spin ephemeris in order to express $E$ as a function of $T$.

Once known, $\bar{P}_{\text{spin}}$ may be used in Equation 6 to determine the average length of the beat period ($\bar{P}_{\text{beat}}$) between $T$ and $T_0$. Thus, the beat epoch relative to $T_0$ is

$$C_{\text{beat}} = \frac{T - T_0}{\bar{P}_{\text{beat}}},$$ \hspace{1cm} \text{(8)}

the decimal portion of which is $\phi_{\text{beat}}$. In our beat-phase calculations, we arbitrarily selected $T_0 = 2449638.3278$ from our cubic spin ephemeris, so at $\phi_{\text{beat}} = 0.0$, the spin phase is 0.0 and the orbital phase is 0.86.

3.3. Variations in Eclipse O–C

3.3.1. Periodicity

In a conference abstract, Geckeler & Staubert (1999) first reported the discovery of a 200-second O–C shift in V1432 Aql’s eclipse timings. We followed up on this periodicity by performing an O–C analysis on all eclipse timings listed in Table I. We calculated both
the O–C timing residual and $C_{\text{beat}}$ for each eclipse and then used the analysis-of-variation (ANOVA) technique \cite{Schwarzenberg-Czerny1996} to generate several periodograms, with $C_{\text{beat}}$ serving as the abscissa.

The first periodogram used all of the eclipse timings reported in Table 1 and it showed a moderately strong signal at $1.00 \pm 0.02$ cycles per beat period, with the folded eclipse timings exhibiting a sawtooth waveform. We then recalculated the power spectrum after adding previously published optical eclipse timings by Patterson et al. (1995), Watson et al. (1995), and Bonnardeau (2012) to the dataset.

Fig. 3.— From left to right: the power spectrum of the timing residuals of the combined dataset described in Section 3.3 and the waveform of the combined dataset when phased at the beat period. Black data points represent our data as listed in Table 1, while gray data points indicate previously published data as explained in Section 3.3.

The combined dataset consists of 159 measurements spanning a total of 138 beat cycles. The strongest signal is at the beat period ($1.001 \pm 0.002$ cycles per beat period), and its waveform consists of an abrupt 240-second shift in the timing variations near $\phi_{\text{beat}} \sim 0.5$, which is when the residual eclipse flux is strongest, as discussed in Section 3.4.1. Both the periodogram and waveform are shown in Figure 3. Between $0.5 < \phi_{\text{beat}} < 0.85$, the eclipses occur $\sim 120$ seconds early, but after $\phi_{\text{beat}} \sim 0.85$, the eclipses begin occurring later, and by $\phi_{\text{beat}} \sim 1.0$, the eclipses are occurring $\sim 120$ seconds late. Although the 240-second O–C jump at $\phi_{\text{beat}} \sim 0.5$ is the most obvious feature in the O–C plot, there is a 120-second jump towards earlier eclipses at $\phi_{\text{beat}} \sim 0.0$. Considering the gradual eclipses, the WD must be surrounded by an extended emission region, so these eclipse timings do not track the actual position of the WD.
3.3.2. Description of Model

Given the asynchronous nature of the system and the ability of the stream to travel most of the way around the WD (Mukai et al. 2003), we hypothesize that cyclical changes in the location of the threading region are responsible for the O–C variation. In an asynchronous system, the position of the threading region can vary because the WD rotates with respect to the accretion stream, causing the amount of magnetic pressure at a given point along the stream to vary during the beat period. Threading occurs when the magnetic pressure \((\propto r^{-6})\) balances the stream’s ram pressure \((\propto v^2)\). For a magnetic dipole, the magnetic flux density \(B\) has a radial dependence of \(\propto r^{-3}\), but with an additional dependence on the magnetic latitude; the magnetic pressure will be even greater by a factor of 4 near a magnetic pole as opposed to the magnetic equator. An additional consideration is that the stream’s diameter is large enough that the magnetic pressure varies appreciably across the stream’s cross section (Mukai 1988).

One of us (KM) modeled this scenario using a program which predicts eclipsing ingress and egress of a point given its \(x, y,\) and \(z\) coordinates within the corotating frame of the binary. The physical parameters used in the program are \(P_{\text{orb}} = 3.365664\) h (measured), \(M_{\text{WD}} = 0.88M_\odot\), \(M_{\text{donor}} = 0.31M_\odot\), \(R_{\text{donor}} = 2.47 \times 10^{10}\) cm, \(i = 76.8^\circ\), and binary separation \(a = 8.4 \times 10^{10}\) cm. The code treats the donor star as a sphere for simplicity, but since we do not attempt to comprehensively model the system in this paper, the errors introduced by this approximation should be minimal. For instance, as a result of this approximation, we had to decrease \(i\) by 0.9° compared to the value from Mukai et al. (2003) in order to reproduce the observed eclipse length.

We first calculated the ballistic trajectory of the accretion stream and arbitrarily selected four candidate threading regions along the stream (P1, P2, P3, and P4) under the assumption that the stream will follow its ballistic trajectory until captured by the magnetic field (Mukai 1988). The eclipse-prediction program then returned the phases of ingress and egress for each of the four points given their \(x\) and \(y\) coordinates within the corotating frame of the binary. We selected these four points arbitrarily in order to demonstrate the effects that a changing threading region would have on eclipse O–C timings; we do not claim that threading necessarily occurs at these positions or that this process is confined to a discrete point in the \(x, y\) plane. Figure 4 shows a schematic diagram of this model.

Our model requires that the ballistic stream be able to travel to P4 and beyond, an assumption which is firmly grounded in previous studies of V1432 Aql and other asynchronous polars (see Section 1). The inefficient threading in asynchronous systems could be indicative of a relatively weak magnetic field or a high accretion rate. For example, Schwarz et al. (2005) found that if the accretion rate in BY Cam were 10-20 times higher than normal
accretion rates in polars, the stream could punch deeper into the WD’s magnetosphere due to the increased ram pressure. Although it is at least plausible that the asynchronism itself causes the inefficient threading, it is not immediately apparent why this would be.

Once threading occurs, the captured material will follow the WD’s magnetic field lines until it accretes onto the WD. To simulate the magnetically channeled portion of the stream, we assumed that captured material travels in a straight line in the $x, y$ plane from the threading region to the WD while curving in the $z$ direction, where $z$ is the elevation above or below the $x, y$ plane. This is another simplification since the magnetic portion of the stream might be curved in the $x, y$ plane, but we expect that the approximation is reasonable. Since the magnetic field lines will lift the captured material out of the orbital plane, we calculated the $x, y$ coordinates of the midpoint between each threading region and the WD and computed its ingress and egress phases at several different values of $z$. Figure 4 shows a schematic diagram of the system used in our model. We reiterate that this is not a comprehensive model, but as we explain shortly, it is sufficiently robust to offer an explanation for the observed O–C variations.

3.3.3. Orientation of the Poles

Before this model is applied to the observations, it is helpful to determine the orientations of the poles at different points in the beat cycle. Since $i \neq 90^\circ$, one hemisphere of the WD is preferentially tilted toward Earth, and we refer to the magnetic pole in that
hemisphere as the upper pole. The lower pole is the magnetic pole in the hemisphere which is less favorably viewed from Earth. In isolation, our observations do not unambiguously distinguish between these two poles, but since the midpoint of the spin minimum (i.e., spin phase 0.0) corresponds with the transit of the accretion region across the meridian of the WD (e.g., Staubert et al. 2003), we can estimate when the poles face the donor star. When $\phi_{\text{beat}} \sim 0.15$, the spin minimum coincides with the orbital eclipse, so one of the poles is approximately oriented towards the secondary at that beat phase. Likewise, at $\phi_{\text{beat}} \sim 0.65$, the spin minimum occurs at an orbital phase of $\sim 0.5$, indicating that a pole is roughly facing the P2 region, with the other facing the secondary. But the question remains: Is this the upper pole, or the lower one?

Mukai et al. (2003) relied upon X-ray observations of eclipse ingresses and egresses to differentiate between the upper and lower poles (see their Figure 15 and the accompanying text). The authors took advantage of the fact that as the system moves through its beat cycle, the accretion spots for each pole will appear to move across the WD’s surface. Eventually, each pole will rotate out of view, resulting in a jump in either the ingress or egress timings, depending on which pole has disappeared. Their model predicts that when the upper pole is aimed in the general direction of P4, the X-ray egresses will undergo a shift to later phases as the upper polecap rotates behind the left limb of the WD as seen at egress in their Figure 15. Likewise, the disappearance of the lower pole behind the left limb at the phase of ingress results in a shift toward later phases in the ingress timings. Based on data in Table 5 of Mukai et al. (2003), the egress jump occurs near $\phi_{\text{beat}} \sim 0.9$, so at that beat phase, the upper pole should be pointed toward the P3-P4 region. The egress jump is more distinct than the ingress jump, so we base our identification of the poles on the egress jump only.

Our identification of the upper and lower poles is an inference and should not be viewed as a definitive claim. For our method to be reliable, it would be necessary for the accretion geometry to repeat itself almost perfectly in both 1998 (when Mukai et al. (2003) observed) and 2012, when we first observed V1432 Aql. Even though the accretion geometry does seem to repeat itself on a timescale of two decades (see, e.g., Section 3.5), this may not always be the case, as is evidenced by the unusual behavior of the spin minimum in 2002 (Boyd et al. 2014). If the accretion rate during our observations was different than it was in 1998, there would be changes in the location and size of the X-ray-emitting accretion regions (Mukai 1988). Moreover, Mukai et al. (2003) cautioned that their model was a simplification because the accretion geometry was poorly constrained. For example, they noted that their model did not account for the offset between the accretion region and the corresponding magnetic pole.

If the upper pole is aimed towards P3-P4 near $\phi_{\text{beat}} \sim 0.9$, then the upper pole would
face the donor at $\phi_{\text{beat}} \approx 0.65$ since the WD appears to rotate clockwise as seen from the donor. Thus, the lower pole is likely pointed in the general direction of the donor star near $\phi_{\text{beat}} \approx 0.15$. We provide a sketch of the system in Figure 5 which shows the inferred positions of the polecaps throughout the beat cycle.

![Sketch of system](image)

Fig. 5.— *A sketch indicating the general positions of the accretion spots at different beat phases as seen from the donor star. The black crosses represent accretion spots visible from the donor, and the vertical line is the WD’s spin axis. Section 3.3.3 explains how we inferred the positions of the two magnetic poles.*

3.3.4. Application of Model

Even though the four P points were arbitrarily selected, the results of the eclipse-prediction program provide testable predictions concerning the O–C variations. In our model, the emission from the accretion curtain and the threading region result in a moving centroid which is responsible for an O–C shift with a half-amplitude of about ±120 seconds. When the centroid of emission is in the $+y$ region in Figure 4, the O–C would be positive, and if it were in the $-y$ half of the plot, the O–C would be negative. Eclipses of point sources at P1, P2, P3, and P4 would result in O–C values of 289 seconds, 204 seconds, 0 seconds, and −533 seconds, respectively. As for the midpoints between each point and the WD, the O–C values would be 122 seconds, 103 seconds, 0 seconds, and −289 seconds for the P1, P2, P3, and P4 midpoints, respectively. The O–C values for the midpoints have a negligible dependence on the height above the orbital plane (provided that the secondary can still eclipse that point). Upon comparing these calculated values with the actual observations, it is clear that the actual O–C timings are inconsistent with a centroid near P1, P2, and P4. However, centroids near the midpoints for P1, P2, and P3 would be consistent with the
O−C timings.

It makes sense to expect the centroid of the emission region to have a less dramatic O−C value than the candidate threading points. Because we expect that the magnetically-channeled part of the stream travels from the threading region to the WD, the light from this accretion curtain would shift the projected centroid of emission towards the WD. In addition, since the threading region likely subtends a wide azimuthal range, it would limit the ability of the projected centroid to move very far from the WD. With these considerations in mind, the consistency of the theoretical O−C values for the P1, P2, and P3 midpoints with the observed O−C variations indicates that our model offers a plausible explanation of the O−C timings.

The sudden jump to early eclipses near $\phi_{\text{beat}} \sim 0.5$ occurs when the inferred orientation of the lower pole is toward the general direction of P3-P4. We surmise that the increased magnetic pressure on that part of the stream is able to balance the decreasing ram pressure, resulting in a luminous threading region. Since the P3-P4 vicinity is in the $-y$ half of Figure 4, an emission region there would result in an early ingress. In all likelihood, the centroid of that threading region does not approach P4 or its midpoint because the theoretical O−C values do not agree with the observed values. However, a centroid closer to P3 would result in a less-early eclipse which would be more consistent with the observations.

As the WD slowly rotates clockwise in Figure 4, the eclipse of the threading region and magnetically-channeled stream would gradually shift to later phases. Half a beat cycle after the O−C jump at $\phi_{\text{beat}} \sim 0.5$, the lower pole would be oriented in the general direction of P2 and the upper pole towards P4. As the upper pole’s magnetic pressure increases on the stream in the P3-P4 vicinity, a new threading region would form there, producing the O−C jump observed near $\phi_{\text{beat}} \sim 0.0$. In short, our model predicts the two distinct O−C jumps and explains why they are from late eclipses to earlier eclipses.

Our observations provide direct evidence of the brief, simultaneous presence of two separate emission regions as the system undergoes its O−C jump near $\phi_{\text{beat}} \sim 0.5$. During one beat cycle in July 2014, the time of minimum eclipse flux on JD 2456842 had an O−C of $\sim 140$ seconds, but on the very next night, there were two distinct minima within the same eclipse. Separated by a prominent increase in flux, one minimum had an O−C of $-80$ seconds, while the other had an O−C of 240 seconds. Assuming a WD eclipse duration of 700 seconds (Mukai et al. 2003) centered upon orbital phase 0.0, the optical eclipse on the first night commenced when the donor occulted the WD, implying a lack of emission in the $-y$ half of the plot in Figure 4. However, the egress of that eclipse continued well after the reappearance of the WD, as one would expect if there were an emission region in the $+y$ area; indeed, a centroid of emission near the P1 midpoint would account for the observed
O–C value. On the ensuing night, by contrast, the eclipse began before the disappearance of the WD, and ended almost exactly when the WD reappeared. The implication of these two light curves is that within a 24-hour span between $\phi_{\text{beat}} \sim 0.47 - 0.48$, an emission region shifted from the $+y$ half of the plot to the $-y$ region. We plot these two light curves in Figure 6.

Fig. 6.— Two eclipses observed on consecutive nights with the 80-cm Krizmanich Telescope. Note the different vertical scale for the two panels. The vertical dashed lines indicate the expected phases of the WD’s ingress and egress. On the first night (JD 2456842, Panel A), the eclipse is very deep and begins with the WD’s disappearance, but on the second night (JD 2456843, Panel B), the eclipse starts before the occultation of the WD. These light curves are consistent with the appearance of a new threading region near P3-P4 in our model, indicating that this process requires less than 24 hours to take place.

Our hypothesis that the location of the threading radius is variable has implications for previous works. In particular, Geckeler & Staubert (1997) and Staubert et al. (2003) used the timing residuals of the spin minima to track the accretion spot as it traced an ellipse around one of the magnetic poles. One of their assumptions was that the threading radius is constant, but this is inconsistent with the conclusions we infer from our observations and model of the system. A variable threading radius would change the size and shape of the path of the accretion spot (Mukai 1988)—and therefore, of the waveform of the spin minima timings used in those studies to constrain the accretion geometry.

### 3.4. Variations in the Residual Mid-Eclipse Flux

#### 3.4.1. Periodicity

The WD is invisible during eclipse, leaving two possible causes for the variation in residual eclipse flux: the donor star and the accretion stream. The magnetic field lines of the
WD can carry captured material above the plane of the system, so depending on projection effects, it is conceivable that some of the accretion flow could remain visible throughout the WD’s eclipse. Therefore, as the accretion flow threads onto different magnetic field lines throughout the beat period, the resulting variations in the accretion flow’s trajectory could cause the residual eclipse flux to vary as a function of $\phi_{\text{beat}}$.

After we calculated the beat cycle count ($C_{\text{beat}}$) for each eclipse observation, we generated a power spectrum using the ANOVA method with $C_{\text{beat}}$ as the abscissa and the minimum magnitude as the ordinate. For this particular periodogram, we used only the 71 eclipses observed with the 28-cm Notre Dame telescope due to the difficulty of combining unfiltered data obtained with different equipment. There is only one strong signal in the power spectrum, and it has a frequency of $0.998 \pm 0.012$ cycles per beat period. Figure 7 shows both the periodogram and the corresponding phase plot, with two unequal maxima per beat cycle.

![Figure 7](image.png)

**Fig. 7.—** The power spectrum of the residual flux and a phase plot showing the waveform of the signal at the beat period. Spanning 11.8 beat cycles, these plots use only the observations made with the 28-cm Notre Dame telescope. The double-wave sinusoid in the phase plot is meant to assist with visualizing the data and does not represent an actual theoretical model of the system.

While a double-wave sinusoid provides an excellent overall fit to the residual-flux variations, the observed mid-eclipse magnitude deviated strongly from the double sinusoid on several occasions. Most notably, on JD 2456842, the system plummeted to $V \sim 17.8$ during an eclipse ($\phi_{\text{beat}} = 0.469$) near the expected time of maximum residual flux. But just 24 hours later, the mid-eclipse magnitude had surged to $V \sim 16.2$ ($\phi_{\text{beat}} = 0.485$), which was the approximate brightness predicted by the double-sinusoid fit. These light curves were shown in Figure 6. Interestingly, during a different beat cycle in 2013, we observed the mid-eclipse magnitude to be $V \sim 17$ at $\phi_{\text{beat}} = 0.485$, suggesting that the residual flux might be
persistently lower near this beat phase. Unfortunately, gaps in our data make it impossible to ascertain whether the residual eclipse flux always drops at this particular beat phase.

3.4.2. Application of Model

We propose that the overall variation in mid-eclipse flux is the signature of an accretion curtain whose vertical extent varies as a function of the threading radius. When the threading region is farther from the WD, the stream can couple onto magnetic field lines which achieve such a high altitude above the orbital plane that the donor star cannot fully eclipse them. By contrast, when the threading region is closer to the WD, the corresponding magnetic field lines are more compact, producing a smaller accretion curtain which the donor occults more fully. The schematic diagram in Figure 8 offers a visualization of this scenario.

![Fig. 8. — Two schematic diagrams providing a simplified illustration of our explanation for the residual flux variations at mideclipse. In both panels, the captured material travels in both directions along an illustrative magnetic field line. The secondary is the gray sphere eclipsing the WD, and the threading point is shown as a large +. The inclination of the magnetic axis with respect to the rotational axis was arbitrarily chosen as 30°. The portion of the magnetic stream which travels upward and which is visible at mideclipse is highlighted. The threading point in Panel A is near P4, and its threading radius is 3.6 times larger than that of the threading point in Panel B, when the threading point is near the stream’s closest approach to the WD.](image)

While it is conceivable that the residual flux variation is caused by material within the orbital plane, the available evidence disfavors this possibility. In particular, Schmidt & Stockman [2001] saw no diminution in the strength of high-excitation UV emission lines during an eclipse with considerable residual flux at $\phi_{\text{beat}} = 0.58$. If these emission lines originated
within the orbital plane, they would have faded during the eclipse. Furthermore, if the source of the residual flux were in the orbital plane, the eclipse width would likely correlate with the mid-eclipse magnitude. The eclipses with high levels of residual flux would be long, while the deeper eclipses would be short. We do not see this pattern in our data, and Figure 12 in Boyd et al. (2014) does not show such a correlation, either.

Our model from Section 3.3.2 predicts that the threading radius will vary by a factor of \( \sim 3.6 \) between P4 and the stream’s point of closest approach to the WD. (We reiterate that since these points are meant to be illustrative, this is not necessarily the actual variation in the threading radius.) The point is that at P3, threading would take place significantly deeper in the WD’s magnetosphere than it would at P4. Moreover, since the predicted threading radius would be largest near an O−C jump, this hypothesis predicts that the amount of residual flux would be greatest near those jumps and lowest between them, as is observed in a comparison of Figures 3 and 7. In the case of a magnetic stream originating from a threading region between P2-P4, the midpoint of the stream would be visible if it achieves a minimum altitude of \( z \sim 0.08a \) above the orbital plane, where \( a \) is the binary separation. At P4, this is only one-quarter the predicted threading radius, but at P2 and P3, this is three-quarters of the predicted threading radius.

This hypothesis also explains why some spectra of V1432 Aql during mid-eclipse show intense emission lines (e.g. Watson et al. 1995; Schmidt & Stockman 2001), while others show only weak emission (e.g. Patterson et al. 1995). For each of these previously published spectroscopic observations, we calculated \( \phi_{\text{beat}} \) and found that the ones showing strong emission lines were obtained when the predicted residual flux was near one of its maxima in Figure 7. By contrast, the spectra containing weak emission were obtained when the expected residual flux was approaching one of its minima. If our hypothesis is correct, then the variation in the emission lines is simply the result of the changing visibility of the accretion curtain during eclipse. Watson et al. (1995) floated a somewhat related scenario to account for the presence of emission lines throughout the eclipse, but they disfavored this possibility largely because of the apparent residual flux at X-ray wavelengths.

An excellent way to test our theory would be to obtain Doppler tomograms near the times of maximum and minimum residual eclipse flux. Schwarz et al. (2005) showed that this technique is capable of revealing the azimuthal extent of the accretion curtain in BY Cam, and it would likely prove to be equally effective with V1432 Aql.

We do not have enough data to consider why the residual flux can vary by as much as \( \sim 1.5 \) mag in one day near the expected time of maximum residual flux. Knowing whether the residual flux is always low near \( \phi_{\text{beat}} = 0.47 \) would be a necessary first step in this analysis.
3.5. The Dependence of the Spin Modulation on Beat Phase

As the WD slowly spins with respect to the secondary, the accretion stream will couple to different magnetic field lines, meaning that the spin modulation will gradually change throughout the beat cycle. To explore this variation, we constructed non-overlapping, binned phase plots of the spin modulation in ten equal segments of the beat cycle (e.g., between $0.00 < \phi_{\text{beat}} < 0.10$). As with the residual-eclipse-flux measurements, we used only the data obtained with the Notre Dame 28-cm telescope in order to avoid errors caused by the different unfiltered spectral responses of multiple telescope-CCD combinations. In an effort to prevent eclipse observations from contaminating the spin modulation, we excluded all observations obtained between orbital phases 0.94 and 1.06. We then calculated the beat phase for all remaining observations and used only those observations which fell into the desired segment of the beat cycle. We used a bin width of $0.01P_{\text{spin}}$, and we did not calculate bins if they consisted of fewer than 5 individual observations.

Figure 9 shows these ten phase plots, and several features are particularly striking. For example, the spin minimum near spin phase 0.0 is highly variable. Conspicuous between $0.5 < \phi_{\text{beat}} < 1.0$, it becomes feeble and ill-defined for most of the other half of the beat cycle. Sometimes, the spin minimum is quite smooth and symmetric, as it is between $0.7 < \phi_{\text{beat}} < 0.8$, but it is highly asymmetric in other parts of the beat cycle, such as $0.5 < \phi_{\text{beat}} < 0.6$. Additionally, there is a marked difference between the phase plots immediately before and after the O–C jump, as one would expect if the O–C jump marks a drastic change in the accretion geometry.

There is also a stable photometric maximum near spin phase $\sim 0.6$ which is visible for most of the beat cycle, though its strength is quite variable. We refer to this feature as the primary spin maximum, but it is not as prominent as the spin minimum. Its behavior is unremarkable.

Interestingly, there is another, much stronger photometric maximum which is visible only between $0.0 < \phi_{\text{beat}} < 0.5$. Since this feature shares the WD’s spin period, we refer to it as the second spin maximum. The second spin maximum can be exceptionally prominent in photometry, attaining a peak brightness of $V\sim14.1$ in several of our light curves—which is the brightest that we have observed V1432 Aql to be. When visible, the second spin maximum is present at spin phase $\sim 0.4$, preceding the primary spin maximum by $\sim0.2$ phase units. It begins to emerge near $\phi_{\text{beat}} \sim 0.0$, and gradually strengthens until it peaks between $0.2 < \phi_{\text{beat}} < 0.3$. It then weakens as $\phi_{\text{beat}}$ approaches 0.5, and after the O–C jump near $\phi_{\text{beat}} \sim 0.5$, the second spin maximum is replaced by a dip in the light curve.

Although the second spin maximum consistently appears between $0.0 < \phi_{\text{beat}} < 0.5$,
Fig. 9.—Binned phase plots of the spin modulation at different beat phases, with each bin representing 0.01 spin cycles. Gaps in the light curves are due to eclipses. The second spin maximum ($\phi_{\text{spin}} \sim 0.4$) is strongest in panel C.
it vanished in a matter of hours on JD 2456842 ($\phi_{\text{beat}} \sim 0.47$), only to reappear the next night. On the first night, our observations covered two spin cycles, and while the second spin maximum was obvious in the first, it had disappeared by the second. Just 24 hours later, it was again visible in two successive spin cycles. This unexpected behavior coincides with the approximate beat phase at which we would expect the dominant threading region to shift to the P3-P4 region in our model. Nevertheless, our lack of observations near this beat phase precludes a more rigorous examination of this particular variation.

The second spin maximum is very apparent in some previously published light curves of V1432 Aql. For example, Watson et al. (1995) presented light curves of V1432 Aql obtained in 1993 which showcase the gradual growth of the second spin maximum (see Panels B-G of their Figure 2). Using our method of determining the beat phase, we extrapolate a beat phase of 0.96 for the light curve shown in their Panel B and a beat phase of 0.12 for the light curve in their Panel G. The increasing strength of the second spin maximum in their light curves agrees with the behavior that we observed at those beat phases (see our Figure 9). Furthermore, Figure 1 in Patterson et al. (1995) shows the second spin maximum at the expected beat phases. These considerations suggest that the second spin maximum is a stable, recurring feature in optical photometry of V1432 Aql.

The overall predictability of the second spin maximum does not answer the more fundamental question of what causes it. One possibility is that it is the result of an elevated accretion rate on one pole for half of the beat cycle. The apparent gap between the two spin maxima, therefore, might simply be the consequence of an absorption dip superimposed on the photometric maximum, splitting it into two.

A more interesting scenario is that the second spin maximum could be the optical counterpart to the possible third polecap detected by Rana et al. (2005) in X-ray and polarimetric data. In that study, Rana et al. (2005) detected three distinct maxima in X-ray light curves as well as strong circular polarization at spin phase 0.45, which is the approximate spin phase of the second spin maximum in optical photometry. They also detected circular polarization at spin phases 0.1 and 0.7, which are the spin minimum and the primary spin maximum, respectively. Quite fortuitously, the authors obtained their polarimetric observations within several days of the photometric detection of the second spin maximum by Patterson et al. (1995). Thus, it is reasonable to conclude that the circular polarization feature near spin phase 0.45 corresponds with the second spin maximum, consistent with a third accreting polecap.

The conclusions of Rana et al. (2005), coupled with our identification of a second spin maximum, suggest that V1432 Aql might have at least three accreting polecaps—and therefore, a complex magnetic field. However, the available evidence is inconclusive, and follow-up
polarimetry across the beat cycle could provide less ambiguous evidence concerning the WD’s magnetic field structure.

4. Conclusion

We have presented the results of a two-year photometric study of V1432 Aql’s beat cycle. We have confirmed and analyzed the eclipse O–C variations first reported by Geckeler & Staubert (1999), and we found that the residual mid-eclipse flux is modulated at the system’s beat period. We interpret these variations as evidence that the azimuth and altitude of the threading region both vary appreciably as a function of beat phase. Doppler tomography of the system at different beat phases could reveal any changes in the azimuthal extent of the accretion curtain, thereby providing a direct test of our model of the system.

The residual flux levels also occasionally exhibit seemingly unpredictable short-term variation, especially around the time of the large O–C jump near $\phi_{\text{beat}} \sim 0.47$. In the most striking example of this variation, the residual flux increased by $\sim 1.5$ mag in just a 24-hour span. We have insufficient observations to explain this behavior, and more observations near that part of the beat cycle are needed. Amateur astronomers are ideally suited to undertake such an investigation, especially when one considers that our residual-flux analysis utilized a small telescope and commercially available CCD camera. Observers with larger telescopes could also obtain relatively high-cadence photometry to study whether double-minima eclipses consistently appear when the system is undergoing an O–C jump.

In addition, we report a second photometric spin maximum which appears for only about half of the beat cycle. This phenomenon might be evidence of a complex magnetic field, but a careful polarimetric study of the beat cycle would be necessary to investigate this possibility in additional detail.

We also offer updated ephemerides of the orbital and spin periods. An exponential spin ephemeris models the data as well as a polynomial ephemeris and is consistent with an asymptotic approach of the spin period toward the orbital period. According to the exponential ephemeris, the rate of change of the spin period is proportional to the level of asynchronism in the system.

Finally, while a comprehensive theoretical model of V1432 Aql is beyond the scope of this paper, such an analysis could refine our model and shed additional light on V1432 Aql’s unusual threading mechanisms.
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*This preprint was prepared with the AAS LaTeX macros v5.2.*
Table 1: Observed Times of Minimum Eclipse Flux

| BJD<sup>a</sup> | ±   | Obs. | BJD   | ±   | Obs. | BJD   | ±   | Obs. |
|-----------------|-----|------|-------|-----|------|-------|-----|------|
| 117.75353       | 0.00052 | CL   | 486.71075 | 0.00055 | JU   | 562.57949 | 0.00047 | CL   |
| 121.67928       | 0.00060 | CL   | 486.85131 | 0.00037 | JU   | 565.66515 | 0.00061 | *    |
| 129.67477       | 0.00051 | CL   | 486.85137 | 0.00039 | CL   | 566.64823 | 0.00074 | CL   |
| 129.81416       | 0.00050 | CL   | 487.69245 | 0.00051 | JU   | 573.65832 | 0.00071 | *    |
| 131.49702       | 0.00040 | DB   | 487.69267 | 0.00051 | CL   | 574.63842 | 0.00048 | CL   |
| 132.47891       | 0.00047 | DB   | 488.81394 | 0.00044 | MC   | 575.61915 | 0.00043 | †    |
| 133.46253       | 0.00073 | DB   | 490.77791 | 0.00051 | CL   | 576.60209 | 0.00040 | CL   |
| 134.44280       | 0.00055 | DB   | 503.68090 | 0.00072 | CL   | 577.58372 | 0.00037 | ‡    |
| 138.51071       | 0.00058 | DB   | 506.62747 | 0.00099 | CL   | 579.54683 | 0.00033 | CL   |
| 145.80112       | 0.00028 | CL   | 506.76566 | 0.00079 | CL   | 580.52958 | 0.00050 | CL   |
| 162.77047       | 0.00040 | CL   | 508.72899 | 0.00047 | CL   | 593.57155 | 0.00031 | ◊    |
| 175.67046       | 0.00038 | CL   | 510.69232 | 0.00055 | MC   | 594.55383 | 0.00040 | CL   |
| 180.57845       | 0.00040 | CL   | 515.60006 | 0.00051 | CL   | 600.58074 | 0.00045 | MC   |
| 180.71866       | 0.00044 | CL   | 528.64310 | 0.00038 | CL   | 787.79453 | 0.00054 | CL   |
| 181.70019       | 0.00043 | CL   | 529.62302 | 0.00021 | CL   | 799.85494 | 0.00046 | CL   |
| 182.68111       | 0.00044 | CL   | 529.76522 | 0.00039 | CL   | 801.81835 | 0.00058 | CL   |
| 194.60329       | 0.00044 | CL   | 530.60621 | 0.00049 | CL   | 813.73962 | 0.00081 | CL   |
| 428.79371       | 0.00028 | CL   | 531.58818 | 0.00026 | CL   | 814.72096 | 0.00043 | CL   |
| 431.87845       | 0.00056 | CL   | 534.67350 | 0.00047 | CL   | 815.70217 | 0.00053 | CL   |
| 447.86614       | 0.00030 | CL   | 538.59720 | 0.00041 | CL   | 815.84330 | 0.00122 | CL   |
| 451.79367       | 0.00039 | CL   | 539.57898 | 0.00030 | CL   | 822.85472 | 0.00077 | CL   |
| 460.76848       | 0.00041 | CL   | 539.71935 | 0.00049 | CL   | 842.76944 | 0.00026 | CL   |
| 462.73258       | 0.00035 | CL   | 540.70141 | 0.00035 | CL   | 843.74858 | 0.00055 | CL   |
| 463.71538       | 0.00066 | CL   | 545.60953 | 0.00054 | CL   | 843.75135 | 0.00087 | JU   |
| 477.73569       | 0.00051 | CL   | 546.59071 | 0.00053 | CL   | 843.75227 | 0.00055 | CL   |
| 484.74745       | 0.00037 | CL   | 548.69482 | 0.00044 | CL   | 847.67489 | 0.00053 | JU   |
| 484.74771       | 0.00071 | JU   | 549.67629 | 0.00038 | CL   | 847.81384 | 0.0015  | JU   |
| 484.88740       | 0.00065 | JU   | 558.65164 | 0.00039 | CL   | 848.65653 | 0.00031 | CL   |
| 485.72961       | 0.00046 | JU   | 559.63391 | 0.00037 | CL   | 849.77851 | 0.00033 | CL   |
| 485.72974       | 0.00041 | CL   | 560.61527 | 0.00056 | CL   |                |        |      |

*: CL, RM
†: RM, RC
‡: RM, RC, KCM, TC

<sup>a</sup>2456000+
Table 2: Observed Times of Spin Minima

| BJD   | ±     | Observer |
|-------|-------|----------|
| 117.8317 | 0.0018 | CL       |
| 119.6569 | 0.0019 | CL       |
| 121.6249 | 0.0021 | CL       |
| 129.7710 | 0.0026 | CL       |
| 131.4553 | 0.0026 | DB       |
| 132.4406 | 0.0031 | DB       |
| 162.6717 | 0.0014 | CL       |
| 175.6023 | 0.0018 | CL       |
| 180.6599 | 0.0021 | CL       |
| 181.6458 | 0.0022 | CL       |
| 182.6293 | 0.0023 | CL       |
| 194.5668 | 0.0021 | CL       |
| 431.8292 | 0.0022 | CL       |
| 484.6855 | 0.0025 | JU       |
| 484.8263 | 0.0022 | JU       |
| 485.6698 | 0.0017 | JU       |
| 485.8085 | 0.0021 | JU       |
| 486.7913 | 0.0024 | JU       |
| 486.7925 | 0.0018 | CL       |
| 487.7752 | 0.0021 | JU       |
| 488.7581 | 0.0029 | MC       |
| 490.7272 | 0.0015 | CL       |
| 528.6812 | 0.0016 | CL       |
| 534.7212 | 0.0018 | CL       |
| 539.6384 | 0.0012 | CL       |
| 540.6261 | 0.0045 | CL       |
| 546.6710 | 0.0019 | CL       |
| 549.6188 | 0.0016 | CL       |
| 558.6081 | 0.0023 | CL       |
| 560.5729 | 0.0023 | CL       |
| 593.6136 | 0.0018 | RM, CL   |
| 594.5978 | 0.0016 | CL       |
| 607.5310 | 0.0019 | RM, RC, KCM, TC, DS |

\(^{a}2456000+\)