β CEPHEI STARS FROM A PHOTOMETRIC POINT OF VIEW

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Abstract. This is an observational review, with an emphasis on photometric data and their interpretation. Two lists are presented, one containing β Cephei stars, and the other, β Cephei suspects. These lists then serve as a basis for discussing such topics as the location of β Cephei stars in the observational and theoretical H-R diagrams, the evolutionary state of these stars, the period-luminosity and period-luminosity-color relations, and observational identification of pulsation modes. The paper also includes references to recent work connected with the theoretical discovery that an opacity mechanism is responsible for the excitation of β Cephei-star pulsations. Finally, observational programs for verifying the consequences of this discovery are suggested.

Key words: β Cephei stars – Photometry

1. Introduction

1.1. Background

The β Cephei variables are a group of early B stars which exhibit short-period variations of brightness, radial velocity, and line profiles. The periods range from about two to seven hours, and are too short to be accounted for by purely geometric effects such as rotation or binary motion. Moreover, the amplitudes of the light curves increase toward shorter wavelengths, indicating a temperature variation. Clearly, the observed variations are due to pulsations. Since they persist over extended intervals of time, the pulsations must be sustained by a mechanism of vibrational instability.

Unsuccessful attempts to find an acceptable mechanisms went on since Christy (1966) pointed out that the region of ionization of He^+, responsible for destabilizing Cepheids and RR Lyrae variables, cannot be effective in β Cephei stars. The problem may have finally been solved. The solution involves the usual κ - mechanism (Baker and Kippenhahn 1962), acting in a zone with temperatures of about 200,000 K, where the metal contribution to stellar opacities has been recently revised. As of this writing, the latest and most complete references on the subject are Dziembowski and Pamyatnykh (1992a, 1992b). We shall return to the problem of the pulsation mechanism in Section 6.

The variability of the radial velocity of β Cep, the prototype, was discovered by Frost in 1902 at the Yerkes Observatory. Frost has also determined that the period

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was equal to $4^h\ 34^m$ (Frost 1902, 1906). Several years later, Guthnick (1913) detected light variations with an amplitude of $0\overset{m}{.}05$ and a period equal to that of the radial-velocity variations.

In 1908 Albrecht at the Lick Observatory found that β CMa showed similar radial-velocity variations (Albrecht 1908). During 1930-32 this star was extensively studied by Meyer (1934). He discovered that the star's radial-velocity amplitude varied periodically, the period amounting to 49.1 days. Meyer was able to explain this in terms of interference between two sine curves of slightly different short periods. Similar long-period velocity and brightness amplitude-variations, referred to as a "beat phenomenon", were later discovered in several other β Cephei stars. In the case of β CMa, the reality of the above-mentioned two sine-curve components was supported by the fact that one of the short periods was identical with the period of variation of the widths of spectral lines, discovered by Henroteau (1918). However, it later turned out that more than two short periods were present in most "beat" β Cephei variables. For example, a third short period was found in the light- and radial-velocity variations of β CMa by Shobbrook (1973a). For the sake of illustration, a sample of Meyer's radial-velocity observations of β CMa is displayed in Figure 1, and some of Shobbrook's (1973a) light curves are reproduced in Figure 2.

Since β CMa became the first well-studied member of the group, these variables have also been called "β Canis Majoris stars." However, there is now a general tendency to use the historically correct name "β Cephei stars."

Following Meyer's work, intensive observational research on β Cephei stars was initiated in the fifties by Struve and his co-workers. A summary of the results was presented by Struve (1955b). The progress had somewhat slowed down after Struve's death, and an apparent status-quo was reached in the seventies. At the end of this period, the observational knowledge and the theoretical situation were extensively reviewed by Lesh and Aizenman (1978).

In their review paper, Lesh and Aizenman (1978) discuss observed properties of "confirmed β Cephei stars", that is, those for which the same short period had been found in both light and radial-velocity variations. They list 21 confirmed β Cephei stars, all brighter than about magnitude 6.5. Spectral types of these stars range from B0.5 to B2, and MK luminosity classes, from II-III to IV-V.

As far as variability of the confirmed β Cephei stars is concerned, the range of light variations is usually less than $0\overset{m}{.}1$ in the visible, and the radial-velocity range seldom exceeds 50 km s^{-1}. Exceptions are BW Vul and σ Sco. In the case of BW Vul, the range of light variations is equal to about $0\overset{m}{.}2$ in V, and the radial-velocity range amounts to slightly more than 200 km s^{-1}, while σ Sco shows a radial-velocity range exceeding 100 km s^{-1}.

As we have already mentioned, the variations are often multiperiodic, but singly-periodic variables also occur. Examples of the latter are δ Cet (Jerzykiewicz *et al.* 1988), ξ^1 CMa (Shobbrook 1973b), and HR 6684 = V2052 Oph (Kubiak and Seggewiss 1984). BW Vul also belongs to this category (vander Linden and Sterken

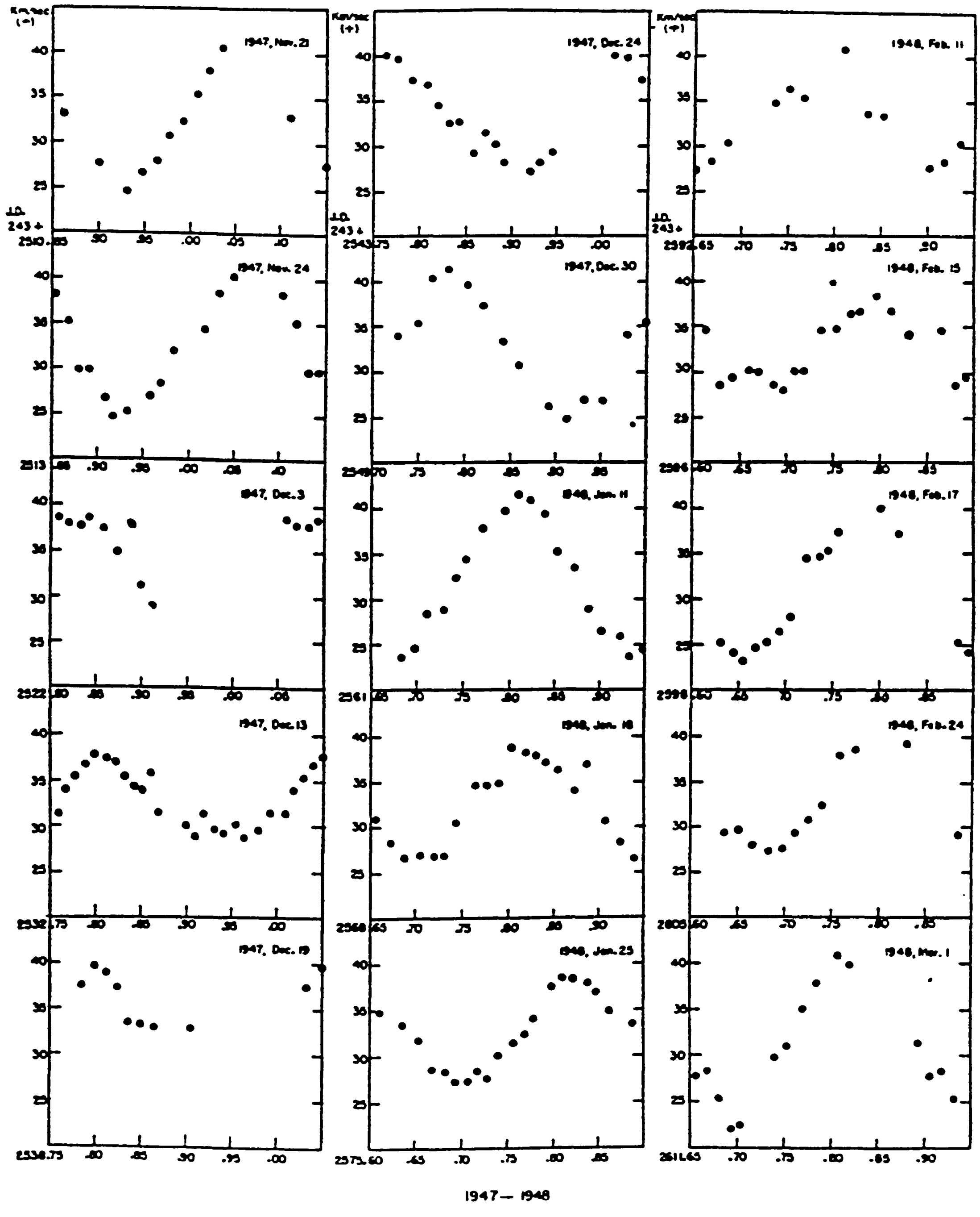

Fig. 1. Meyer's radial-velocity observations of β CMa, reproduced from Struve (1950).

1987).

Among the 15 investigated cases, 13 show line-profile variations. As a rule, the spectral lines are broadest on the descending branch of the radial-velocity curve. Line splitting is sometimes associated with these profile variations.

In addition to the confirmed β Cephei variables, Lesh and Aizenman (1978) list several "β Cephei candidates", that is, early B stars found to show short-period light variations, but lacking confirmation of variability in radial velocity. It should

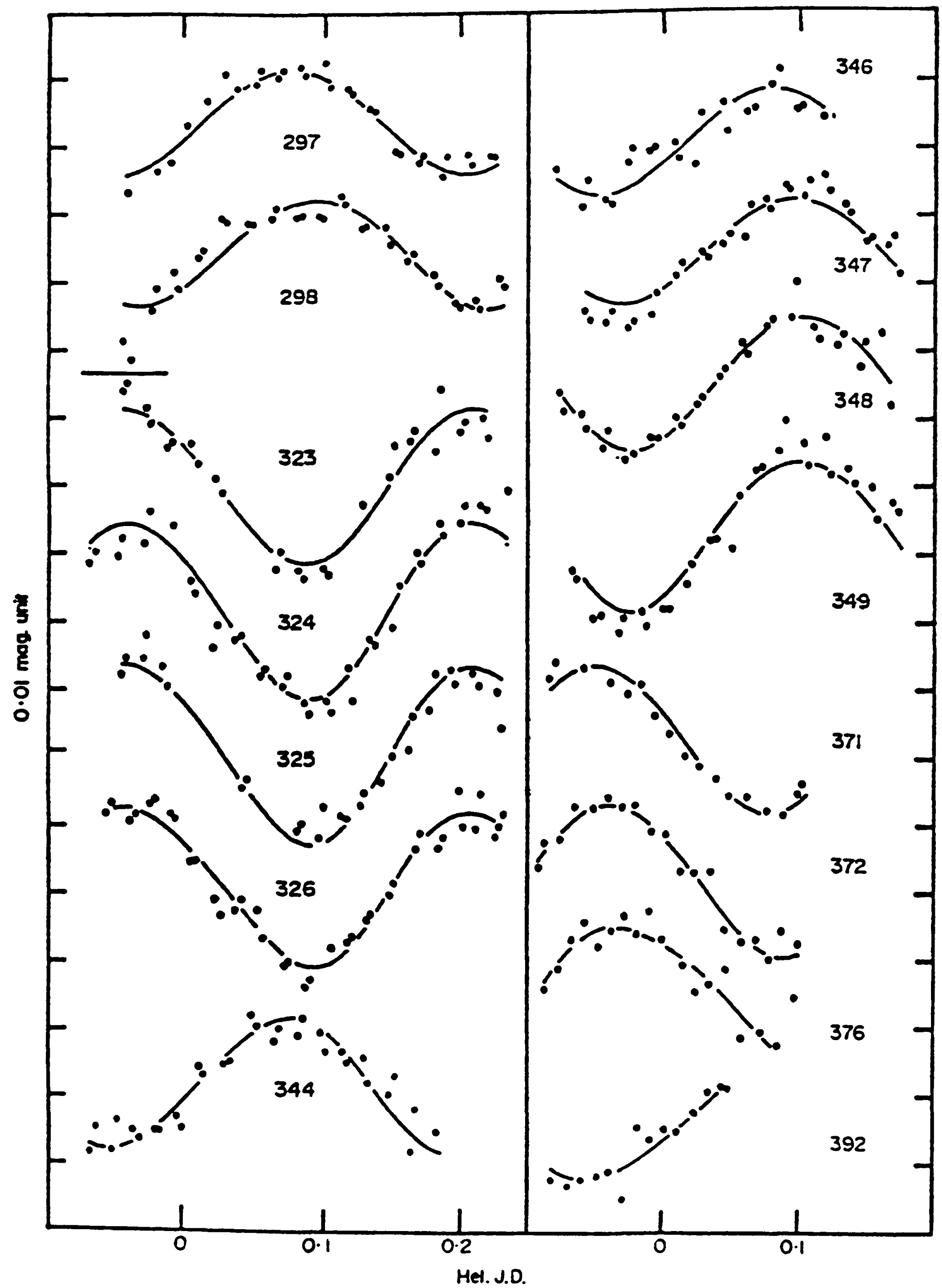

Fig. 2. Shobbrook's (1973a) yellow-filter observations of β CMa (points). The number against each graph is J.D. *minus* 2,441,000. The solid lines were fitted to the data using the three periods mentioned in the text.

be pointed out that a number of objects labeled as β Cephei stars in the literature and, particularly, in the *General Catalogue of Variable Stars* (GCVS, Kholopov *et al.* 1985b, 1985c, 1987a), are merely β Cephei candidates in the sense of the above definition. In the present paper we follow this practice: we consider light variability as a sufficient condition to classify an early B star as a β Cephei variable, provided

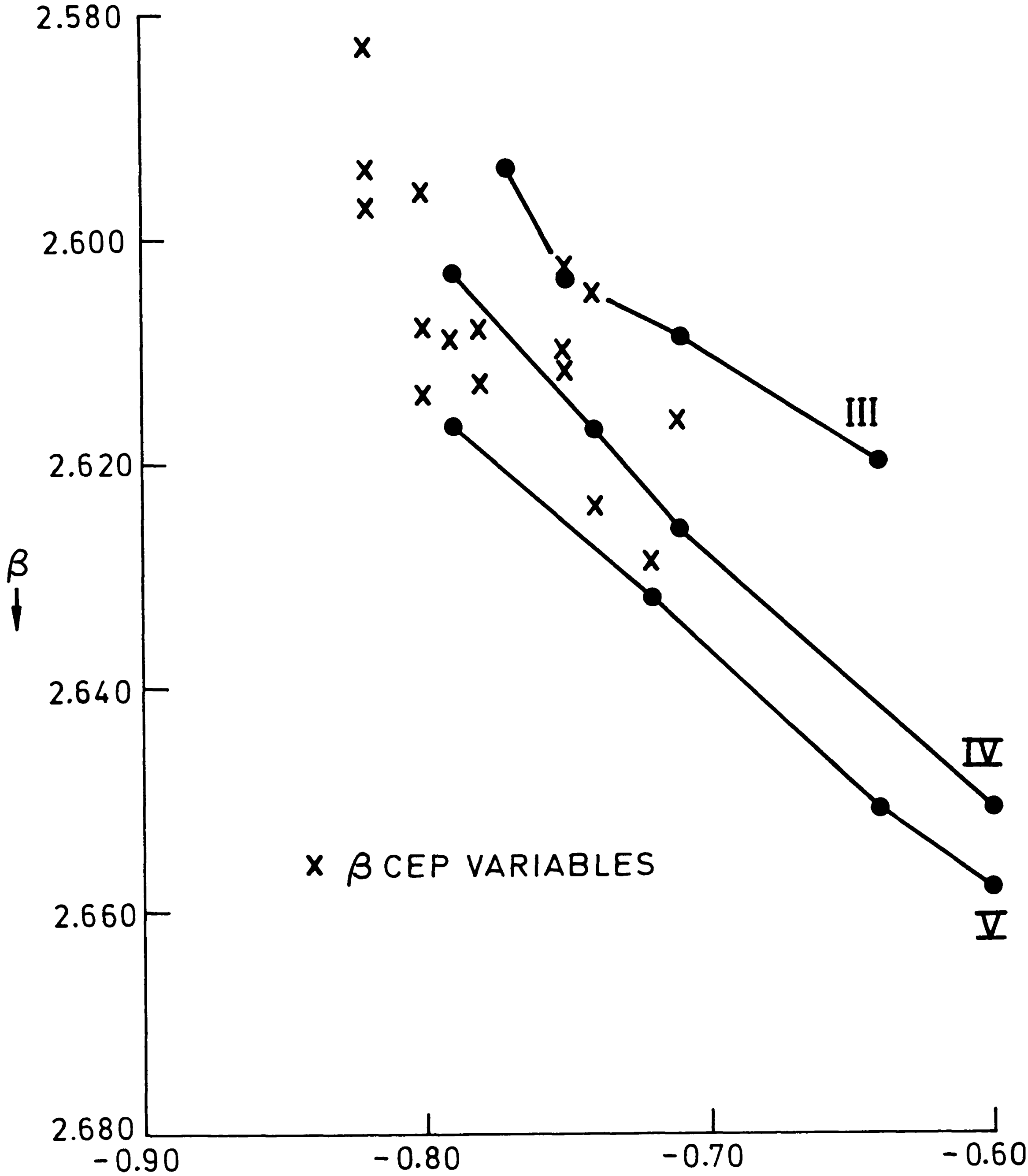

Fig. 3. Location of "confirmed" β Cephei variables in a color-magnitude diagram based on the photometric parameters Q and β, and average sequences for MK luminosity classes III, IV, and V, spectral types from B1 to B2.5 (after Lesh and Aizenman 1973a).

that the period is shorter than about $0\overset{d}{.}3$. This point will be discussed further in the next section.

The small ranges of spectral type and luminosity class occupied by β Cephei stars make them cluster in a small region of the H-R diagram. This region is commonly labeled as the "β Cephei instability strip". Figure 3, reproduced from Lesh and Aizenman (1973a), gives the location of the confirmed β Cephei stars in the observational H-R diagram, defined by the reddening-free photometric parameter

Q of the UBV system (Johnson and Morgan 1953, see also below, Section 3.2), and Crawford's (1966) β index. The former quantity, which is a function of the size of the Balmer jump, measures the effective temperature, while the latter, related to the strength of the Hβ line, is an absolute-magnitude index. In Figure 3, the confirmed β Cephei variables are indicated by crosses, and the mean points for nonvariable stars of spectral types B1 - B2.5 and luminosity classes III, IV, and V are also shown.

1.2. Questions

In addition to the fundamental problem of the pulsation mechanism, there were the following questions to which Lesh and Aizenman (1978) found conflicting answers in the literature:

1. What are, precisely, the high- and low-temperature limits of the β Cephei instability strip?
2. How wide is the instability strip in absolute magnitude?
3. What is the evolutionary state of these stars?
4. What is the frequency of occurrence of β Cephei stars, and is this frequency compatible with their evolutionary state?
5. Do variable and nonvariable stars coexist in the instability strip?
6. Is the dominant pulsation mode the same in all β Cephei stars?
7. What role does rotation play in the β Cephei phenomenon?

Progress achieved in the past fifteen years in answering these questions has been due in part to photoelectric search programs, which considerably increased the number of known variables. In particular, they helped to settle the dispute over the evolutionary state of these stars. Crucial in this respect was the discovery of β Cephei stars in young open clusters. In the first section of the next section, the search programs are reviewed. In the second section of that section, lists of β Cephei stars and β Cephei suspects are presented. In Sections 3 and 4, these lists serve as a basis for discussing answers to the first five questions. Question 6, and the more general problem of observational identification of pulsation modes, are considered in Section 5. Finally Section 6 we devote to the opacity mechanism, mentioned at the beginning of this Introduction.

2. β Cephei stars and β Cephei suspects

2.1. Searches for β Cephei stars

Originally, β Cephei stars were discovered in a more or less random way as by-products of stellar radial-velocity observations. Spectrograms of B stars, often taken for the sake of kinematical studies of the Galaxy, have occasionally indicated radial-velocity variations, characteristic of this type of variability. Subsequent spectrographic follow-up and additional photometric observations could then reveal the true nature of the variables. As we already mentioned in the Introduction, the first β Cephei stars were discovered in this way by Frost (1902), Albrecht (1908), and Guthnick (1913).

The earliest systematic program of photoelectric monitoring of B stars, known or suspected to show radial-velocity variations, has been undertaken by Walker (1952). Since this pioneer work, several programs aimed at discovering β Cephei stars were carried out. Lynds (1959) examined for light variability thirty-six bright stars spectroscopically similar to β Cephei variables. In this case, variable radial velocity was not used as a condition to select a star for observing. Among the twenty-nine stars Lynds (1959) found to vary by $0\overset{m}{.}02$ or more, two he classified as β Cephei variables. The rest comprised eclipsing binaries, ellipsoidal variables, shell stars with irregular light variations, stars having periodic light variations with periods slightly longer than those of β Cephei stars and shorter than would be expected for ellipsoidal variables, and, finally, objects which could not be classified because of irregularities, or because of insufficient data.

The nature of one of the two variables classified by Lynds (1959) as β Cephei stars, V986 Oph, is still a matter of controversy (cf. Jerzykiewicz 1975, Fullerton *et al.* 1985, Cuypers *et al.* 1989).

Another extensive photometric search was carried out by Hill (1967) among a sample of 153 early B stars, mostly members of the nearest associations and galactic clusters. According to Hill (1967), his program yielded 24 new β Cephei variables, including five tentative ones. Unfortunately, for none of Hill's candidates the existence of short-period light variations has been confirmed by other workers. The only exception, V986 Oph, which Hill included among his newly discovered β Cephei stars, was already known to Lynds (1959). In fact, several of Hill's candidates were subsequently shown to be either constant or to exhibit ellipsoidal light variations (Percy 1969, Deupree 1970, Jerzykiewicz 1974, Hill et al. 1976, see also Table II). Hill's conclusion that the limits of spectral type, luminosity class, and period of the β Cephei variables should be considerably extended beyond their "classical" values is, therefore, unfounded.

A program similar to that of Lynds (1959) was undertaken by Jerzykiewicz (1972b, 1992). The outcome was the discovery of one β Cephei star, HR 6684, and two β Cephei suspects (see Table II). The radial-velocity variation of HR 6684 was detected by Pike (1974).

All the above-mentioned search programs were confined to declinations greater than −20°. The first program directed toward discovering β Cephei stars south of this limit was undertaken by Pagel (1956). He examined seven variable radial-velocity B stars for the presence of short periods, and obtained positive results in three cases, all subsequently confirmed.

Shobbrook and his co-workers (Shobbrook *et al.* 1969, Shobbrook and Lomb 1972, Shobbrook 1972) discovered several southern β Cephei variables in the course of observing in detail the stars measured by the Narrabri Stellar Intensity Interferometer (Hanbury Brown 1974).

The first systematic search of the Lynds (1959) type on the southern sky was

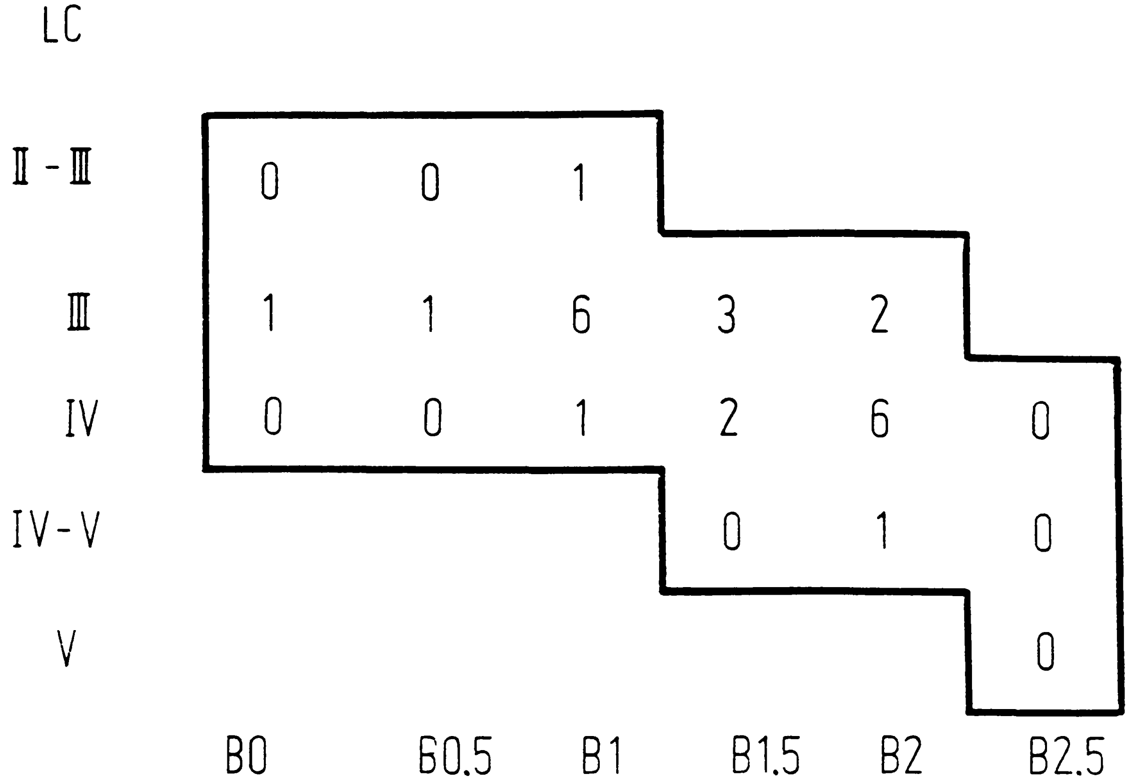

Fig. 4. Distribution of β Cephei stars discovered before 1977 in the spectral type–luminosity class diagram. The number of variables for each MK type is indicated (from Jerzykiewicz and Sterken 1977). The delineated area is the "β Cephei box" discussed in the text.

carried out by Jerzykiewicz and Sterken (Jerzykiewicz and Sterken 1977, Sterken and Jerzykiewicz 1983). These authors compiled a list of all stars south of declination −20°, which appeared in the *Catalogue of Bright Stars* (Hoffleit 1964), and whose position in the luminosity class–spectral type plane was the same, or nearly the same, as that of the β Cephei variables known at that time. Figure 4 shows the boundaries of the region they adopted. This region contained 105 stars which were not previously known as variables, and for which photometric observations could easily be done. Intensive differential observations, programmed to make most likely the discovery of light variations with time scale of about three to seven hours, were carried out during the summer season of 1975, and the winter season of 1977. These observations were complemented with a subsequent detailed photometric and spectrographic investigation (Jerzykiewicz and Sterken 1979, Sterken and Jerzykiewicz 1980). The combined results yielded a total of five new β Cephei stars, among which three had the shortest periods ever found. As a by-product, many new variables of other types were discovered, as well as evidence of nonvariability of a number of early B stars.

The region of the luminosity class-spectral type plane delineated in Figure 4, from which Jerzykiewicz and Sterken (1977) selected stars for observing, was called by them the "β Cephei box". Consequently, Jerzykiewicz and Sterken referred to all their program stars as the "β Cephei box stars". Unfortunately, this has led to confusion, because some authors assumed that these terms are equivalent to

"β Cephei instability strip" and "β Cephei stars" (see, e. g., Sareyan et al. 1980, Hoffleit 1982).

Balona (1977) selected candidates for a photometric search program from lists of Cape photometry where they were noted as suspected variables. He reported eight out of his thirty-one program stars to be previously unrecognized certain or probable β Cephei variables. Four have been confirmed. Most important was Balona's (1977) discovery of three β Cephei stars in NGC 3293, the first variables of this type found in an open cluster.

Jakate (1978) searched for variables another open cluster, NGC 4755 (κ Crucis cluster), and reported finding three β Cephei stars. Jakate (1979a) has also looked for β Cephei stars among 37 southern bright early B stars, and claimed the discovery of two such objects. In addition, he found small-amplitude light variations with periods below one hour in four B2V and B3IV stars, which led him to postulate the existence of a new class of "ultra-short-period B variables" (Jakate 1979b). However, most of these results were subsequently shown to be almost certainly spurious (see, e.g., Balona 1982, Sterken and Jerzykiewicz 1983, Shobbrook 1984).

Jakate and Sterken (1980) stressed the importance of studying β Cephei stars in open clusters and associations, pointing out that the position of these stars in the color-magnitude diagrams would provide an answer to the question about their evolutionary state. Subsequent searches for β Cephei variables in open clusters have been extremely successful. To the already mentioned variables in NGC 3293, found by Balona (1977), Balona and Engelbrecht (1983) added seven new ones. In NGC 6231, Shobbrook (1979a), Balona (1983), and Balona and Engelbrecht (1985a) discovered six β Cephei stars. Recently, Delgado et al. (1984, 1985, 1992) reported results of a photometric search for β Cephei stars in the Northern Hemisphere open clusters NGC 6871, IC 4996, NGC 1502 and NGC 2169.

Among the field B stars, several new β Cephei variables have been found by Waelkens and Rufener (1983a, 1988), Waelkens and Cuypers (1985), Waelkens and Heynderickx (1989), Sterken and Jerzykiewicz (1990a), and Waelkens *et al.* (1991)[1].

2.2. The Lists of β Cephei Stars and β Cephei Suspects

In order to make a list of β Cephei stars, we conducted a literature search for stars that have been classified as members or as possible members of the group on the basis of short periods seen in their light or radial-velocity variations. Among the 138 objects we found, there were 59 with well-documented and convincing evidence for light variations with at least one period shorter than about $0\overset{d}{.}3$. All these stars we consider to be β Cephei variables. They are listed in Table I. The remaining 79 stars, which we shall refer to as β Cephei suspects, are listed in Table II.

[1] For a detailed discussion of some of these results, see Heynderickx (1992)

TABLE I

The β Cephei stars

HR/Cluster	HD/DM	Name	m_V	MK	Period	Reference
39	886*	γ Peg	2.8	B2 IV	$0^d.15175$	dJSa82
779	16582*	δ Cet	4.1	B2 IV	0.16114	JSK88
1072	21803*	KP Per	6.4	B2 IV	0.20175	JJRR81
1463	29248*	ν Eri	4.0	B2 III	0.17351	CG81
2294	44743*	β CMa	2.0	B1 II–III	0.25002	K80a
2387	46328*	ξ^1 CMa	4.3	B1 III	0.20958	Sh73b
2571	50707*	15 EY CMa	4.8	B1 III	0.18456	Sh73b
2648	52918	19 V637 Mon	5.0	B1 IV	0.1912	BCM92
2745	56014	27 EW CMa	4.7	B3 IIIp	0.0918	BR91
—	59864	—	7.6	B1 III	0.238:	SJ90a
2928	61068	PT Pup	5.7	B2 III	0.1612	Sh81
3058	63949	QS Pup	5.8	B1.5 IV	0.1182	JS79
3078	64365*	QU Pup	6.0	B2 IV	0.1927	JS79
3088	64722	V372 Car	5.7	B1.5 IV	0.1154	JS79
3213	68324	IS Vel	5.2	B2 IVn	0.108	JS79
—	78616	KK Vel	6.8	B2 II–III	0.21570	Co82
—	80383	IL Vel	9.1	B2 III	0.18465	Ha79
N3293–11	303068	V400 Car	9.8	B1 V	0.1458	En86
N3293–10	303067	V401 Car	9.6	B0.5 V	0.1689	En86
N3293–16	–57 3500	V403 Car	8.8	B0.5 III–V	0.2506	En86
N3293–65	—	V412 Car:	9.9	—	0.1135	En86
N3293–23	–57 3506	V404 Car	9.2	B1 III	0.1621	En86
N3293–14	–57 3507	V405 Car	9.3	B1 IV	0.1524	En86
N3293–24	–57 3517	V378 Car:	9.2	B1 III	0.2062	En86
N3293–18	–57 3524	V406 Car	9.3	B1 IV	0.1767	En86
N3293–27	92007	V380 Car:	9.0	B0.5 III	0.2273	En86
N3293–5	92024	V381 Car:	9.0	B1 III	0.1773	JS92
4853	111123*	β Cru	1.3	B0.5 III	0.19120	Sh79a
N4755–F	–59 4564	BW Cru	9.0	B2 III	0.2	Sh84
—	112481	V856 Cen	8.3	B2 II–III	0.25454	WH89
5056	116658*	α Vir	1.0	B1 IV	0.1738	SJM86
5132	118716	ϵ Cen	2.3	B1 III	0.16961	Sh72
5267	122451*	β Cen	0.6	B1 III	0.157	KS82
5395	126341*	τ^1 Lup	4.6	B2 IV	0.17735	C87
5469	129056*	α Lup	2.3	B1.5 III	0.25985	Sh79a
5488	129557*	BU Cir	6.1	B2 III	0.12755	vLS85
—	129929	V836 Cen	8.1	B3 V	0.15478	WR83a
5695	136298*	δ Lup	3.2	B1.5 IV	0.1982:	LP88
—	145794	V349 Nor	8.7	B2 II–III	0.15991	WH89
6084	147165*	σ Sco	2.9	B1 III	$0^d.24684$	JS84
—	147985	V348 Nor:	7.9	B1–2 II–III	0.13231	WC85
N6231–253	–41 7706	V945 Sco	9.6	B1:V + B1:V	0.0671	BSh83
N6231–282	–41 7711	—	9.7	B2 V + B2 V	0.1193	BEn85a
N6231–261	–41 7715	V946 Sco	10.0	B2 IV–Vn	0.0988	BSh83

Table I (continued)

HR/Cluster	HD/DM	Name	m_V	MK	Period	Reference
N6231–238	326330	V964 Sco	9.3	B1 V	0.0878	BEn85a
N6231–110	–41 7753	V947 Sco	9.8	B1 V	0.1079	BSh83
N6231–150	326333	V920 Sco	9.6	B1 Vn	0.1012	BSh83
—	156662	V831 Ara:	7.8	B2 III	0.16890	WC85
6453	157056*	θ Oph	3.3	B2 IV	0.14053	Br67
6527	158926*	λ Sco	1.6	B1.5 IV	0.21370	LSh75
6580	160578*	κ Sco	2.4	B1.5 III	0.19987	LSh75
6684	163472*	V2052 Oph	5.8	B2 IV–V	0.13989	KS84
6747	165174	V986 Oph	6.1	B0 IIIn	0.303	CBM89
—	166540	—	8.3	B0.5 IV	0.23299	WVV91
8007	199140*	BW Vul	6.4	B2 III	0.20104	vLS87
8238	205021*	β Cep	3.2	B1 III	0.19049	dJSa82
8640	214993*	12 DD Lac	5.2	B1.5 III	0.19308	JBM84
—	—	PHL 346	11.5	—	0.15231	KW90
8725	216916*	16 EN Lac	5.6	B2 IV	0.16917	PJ88

Remarks

N and I in Column 1 stand for NGC and IC, respectively.

* in Column 2 denotes a confirmed β Cephei star in the sense of Lesh and Aizenman (1978).

: in Column 3 marks a star classified as BCEP: in GCVS or NLVS.

HR 2745: also classified B3 IIIe.

N6231–253, –261, and –110: β Cephei-type variability was discovered by Balona (1983).

N6231–238: was found constant by Balona (1983); short-period variability was subsequently discovered by BEn85a.

N6231–150: variability was discovered by Sh79a.

N6231–282: was classified as a "possible β Cephei variable" by Balona (1983); short-period variability was confirmed by BEn85a.

HD 166540: also classified B1 Ib.

The first three columns of Tables I and II contain catalogue numbers and names of the stars. In columns 1, NGC and IC are abbreviated to N and I, respectively. Two members of the h and χ Persei cluster are identified in Table II by their Oosterhoff (1937) numbers, preceded by Oo. Stellar names in columns 3 are from GCVS and from the four most recent Name-Lists of Variable Stars (NLVS, Kholopov *et al.* 1985a, 1987b, 1989, Kazarovets and Samuś 1990). In a number of cases, Flamsteed numbers are also given. For example, EN Lac is better known in the literature as 16 Lac; both designations are given in column 3 of Table I. The MK classification in columns 5 is from Lesh (1968) and from Hiltner *et al.* (1969) for the HR stars, and from Lesh and Aizenman (1973a) and various other sources for the fainter field stars. For the open-cluster β Cephei stars, the MK types are from Turner *et al.* (1980) for NGC 3293, Perry *et al.* (1976) for NGC 4755, Schild *et al.* (1971) and Levato and Morrell (1983) for NGC 6231. Column 6 of Table I contains periods. For multiperiodic β Cephei stars, the period given is that of the short-period component with greatest light amplitude. For the β Cephei suspects no periods are listed, because in most cases uncertain or conflicting values are reported in the literature. Last columns of Tables I and II contain references. In Table I, a single reference to a recent paper is given for each star. In Table II, the single or the first reference is to a source in which the star has been classified as a β Cephei variable, whereas the second one, to a paper containing evidence to the contrary. In a few cases the source of classification is GCVS or NLVS.

Confirmed β Cephei stars, in the sense of Lesh and Aizenman (1978), that is, those for which the same short period had been found in both light and radial-velocity variations, are indicated in Table I by an asterisk in column 2. There are 24 such stars.

Most stars of Tables I and II are classified in GCVS and NLVS as BCEP or BCEP:, that is, as certain or possible β Cephei stars, respectively. The latter we indicated by a colon in columns 2 of Tables I and II. It can be seen that in a number of cases our Table I stars have been classified as BCEP: by Kholopov and his co-workers. On the other hand, a number of stars which we consider to be merely β Cephei suspects are BCEP according to GCVS or NLVS. In addition, we disregarded all stars, classified in GCVS as BCEPS and called "short period group of β Cephei variables" by Kholopov *et al.* (1985b), because these stars are the spurious "ultra-short-period B variables" of Jakate (1979b), mentioned in the preceding section.

Some of the above-mentioned discrepancies in classification may have been caused by the fact that the GCVS definition of β Cephei variables embraces 53 Persei stars which show light variations, and puts the upper limit of the period at $0\overset{d}{.}6$, instead of $0\overset{d}{.}3$ as we do. Most, however, have apparently resulted from a different evaluation of available data. We believe that our classification is more consistent than that of Kholopov and his co-workers.

3. β Cephei stars in the observational H-R diagrams

In this section, we investigate the location of β Cephei stars in various observational H-R diagrams. We start with the lowest-resolution diagram, that is, the one having spectral type as abscissa and luminosity class as ordinate. Next we treat color-magnitude diagrams most frequently used for early-type stars. This allows us to examine the notion of the β Cephei instability strip and leads to a discussion of the evolutionary state of β Cephei stars that are members of open clusters. We also consider the question of nonvariable stars in the instability strip. In the last section we deal with the β Cephei stars and β Cephei suspects that have been found outside the instability strip.

TABLE II

The β Cephei suspects. The same codes as for Table I were used. Oo are Oosterhoff (1937) numbers in the h and χ Per cluster.

HR/Cluster	HD/DM	Name	m_V	MK	Reference
155	3379	53 AG Psc	5.9	B2.5 IV	Wi54, JS89
—	13051	V351 Per:	8.7	B1 IVe1	Hi67, MM79
—	13494	V352 Per	9.3	B1 III	Hi67
—	13544	V353 Per	8.9	B0.5 IIIn	Hi67
—	13745	V354 Per	7.9	B0 II	Hi67
—	13831	V473 Per:	8.3	B0 IV	GD82, Hi67
—	56 473	V356 Per:	9.1	B0.5 IIIn	Hi67, MM79
—	13866	V357 Per:	7.5	B2 Ib	Hi67, MM79
—	14053	—	8.4	B1 II	Hi67
Oo 963	—	—	10.7	B2 IV	P72
—	14250	V359 Per:	9.0	B1 IV	Hi67
Oo2299	—	—	9.1	B0.5 IV	P72
—	56 589	V360 Per:	9.5	B1 III	Hi67
—	15239	V528 Cas:	8.7	B5 V	PM72
—	15752	V362 Per:	8.7	B0 IIIn	Hi67
—	16429	V482 Cas	7.7	O9.5 III	Hi67
938	19374	53 UW Ari	6.1	B1.5 V	Bo67, S88
1156	23480	23 V971 Tau	4.2	B6 IV	NLVS3
1215	24640	—	5.5	B1.5 V	Jo60
1220	24760	ϵ Per	2.9	B0.5 III	NLVS2, H87
N1502–A	—	—	6.9	B2 IV + B2.5 V:	Hi67
N1502–26	—	—	9.7	B1 V	DAGG92
1350	27396	53 V469 Per	4.8	B4 IV	GCVS3, B87b
1417	28446	1 Cam	5.8	B0 IIIn	J92
1679	33328	λ Eri	4.3	B2 IVn	B77, BCM92
—	34656	—	6.8	O7 IIf	FGB91
1788	35411	η Ori:	3.4	B0.5 Vnn	GCVS2, WL88
1811	35715	ψ Ori	4.6	B1 V	Hi67, J84
1855	36512	υ Ori	4.6	B0 V	BEn85b
—	37776	V901 Ori	7.0	B2 IV	Hi67, TL85
N2169–2	252214	V916 Ori	8.1	B2.5 V	Hi67
N2169–5	252248	V917 Ori	8.8	B2 V:	Hi67
—	43078	LR Gem	8.8	B0.5 III	Hi67, JL81
—	43818	LU Gem	6.9	B0 II	Hi67, P84
—	43837	V963 Ori:	8.4	B2 Ib	P70a
2442	47432	—	6.2	O9.5 II	B77, BEn81
2596	51309	ι CMa	4.4	B3 II	E79a, B77
2603	51630	HH CMa	6.6	B1–2 III:	B77, SJ83
2670	53755	V569 Mon	6.5	B0.5 IVn	B77
2678	53974	FN CMa	5.4	B0.5 III	Hi67, vH73b
2734	55857	GY CMa:	6.1	B0.5 V	vH73a, SJ83
2741	55958	GG CMa	6.6	B2 IV	vH73a
2790	57219	NW Pup	5.1	B2 IVne	vH73a
3117	65575	χ Car	3.5	B3 IVp	E79a

Table II (continued)

HR/Cluster	HD/DM	Name	m_V	MK	Reference
3186	67536	V375 Car:	6.3	B2.5 Vn	JS77, JS79
3440	74071	HW Vel	5.5	B5 V	GCVS3, SvL89
3447	74195	o Vel	3.6	B3 IV	vH72, WR83b
3454	74280	η Hya	4.3	B4 V	E79a
3457	74375	V343 Car:	4.3	B1.5 III	vH73a, B77
3593	77320	IU Vel	6.0	B3 Vne	BBR80, BCM92
3924	85953	—	5.9	B2 V	Ja79a, SJ83
3941	86466	IV Vel	6.1	B3 IV	Ja79a, SJ83
4064	89688	23 RS Sex:	6.7	B2.5 IV	dJ53, J72a
N3293–26	–57 3526	V379 Car	8.5	B1 III	B77, BEn83
—	96446	V430 Car	6.7	B1 IVp	MB91
—	—	V770 Cen:	12.4	B5e	VWPB74
N3766–67	—	V847 Cen:	9.8	B2 Vp	BEn86
4603	104841	θ^2 Cru:	4.7	B2 V	E79a
4656	106490	δ Cru	2.8	B2 IV	E79b, Sh81
4798	109668	α Mus	2.7	B2 IV–V	E79a, B77
N4755–G	–59 4528	BS Cru:	9.8	B0.5 V	Ja78
N4755–IV–18	–59 4542	BT Cru	9.6	B2: V	Ja78, Sh84
N4755–I–05	—	BV Cru	8.6	B0.5 IIIn	Ja78
4897	112078	λ Cru:	4.6	B4 Vn	GCVS1, Sh81
5034	116072	V790 Cen:	6.2	B2.5 Vn	GCVS1
5190	120307	ν Cen	3.4	B2 IV	R77, CBM89
5934	142883	—	5.9	B3 V	Hi67
—	149881	V600 Her:	7.0	B0.5 III	HiDO76, J74
N6231–28	326327	V962 Sco	9.7	B1.5 Ve+shell	BEn85a
N6231–289	–41 7724	—	9.5	B0.5 V	BEn85a
N6231–80	–41 7734	V963 Sco:	10.4	—	BEn85a
6588	160762	ι Her	3.8	B3 IV	CLV87
7318	180968	2 ES Vul:	5.4	B1 IV	Ly59
7600	188439	V819 Cyg:	6.3	B0.5 IIIn	Ly59
7647	189687	25 V1746 Cyg	5.2	B3 IV	PJaM81
N6871–14	—	V1820 Cyg:	10.8	B2–3 IV–III	DAG84
I4996–7	—	V1922 Cyg	10.9	—	DAG85
8105	201819	—	6.5	B0.5 IVn	J92
—	217035	KZ Cep	7.8	B0.2 IV	Hi67, De70
9070	224559	LQ And:	6.5	B4 Ve3n	PJaM81

Remarks

N and I in Column 1 stand for NGC and IC, respectively, while Oo

: in Column 3 marks a star classified as BCEP: in GCVS or NLVS.

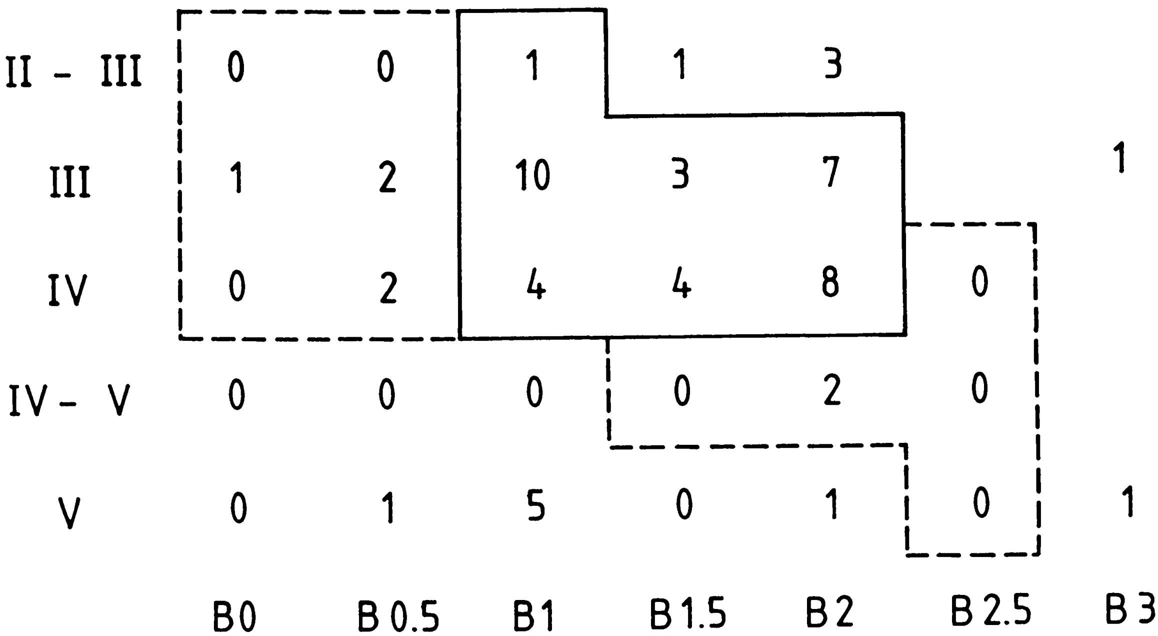

Fig. 5. Updated version of Figure 4, based on the MK types listed in Table I. The area occupied by the ten β Cephei stars known to Struve (1955b) is delineated with solid lines. The "β Cephei box" of Figure 4 is indicated by means of dashed lines.

3.1. The Spectroscopic H-R Diagram

Figure 5 shows the distribution of β Cephei stars in the spectral type–luminosity class plane. Included are all stars of Table I, except NGC 3293-65 = V412 Car and PHL 436 for which no classification on the MK system is available. The ten β Cephei stars known to Struve (1955b) fall within the area delineated with solid lines.

The star at location B0 in Figure 5 is V986 Oph. The B0 type is approximately consistent with the star's color indices (see the next section). On the other hand, both UBV and Geneva color indices of V836 Cen, one of the two B3 stars in the figure, correspond to a spectral type of about B1 (see below, Sections 3.2 and 3.4). The other B3 star in Figure 5 is 27 (EW) CMa, the Be star in which β Cephei light variations have been recently discovered by Balona and Rozowsky (1991). Thus, except for V986 Oph and 27 (EW) CMa, the β Cephei stars are confined to the very narrow spectral-type range from B0.5 to B2. That is, *the photometric search programs failed to significantly extend the spectral type range occupied by β Cephei stars beyond the limits set by Struve (1955b).* However, several β Cephei stars of luminosity class V were recently found, whereas β Cephei stars known to Struve had luminosity classes from IV to II-III. As can be seen from Table I, most of the class V variables are members of NGC 3293 or NGC 6231.

A different conclusion concerning the spectral types at which β Cephei stars occur has been reached by Sareyan et al. (1980), who maintain that almost all searches for β Cephei variables were restricted to the small region of the H-R

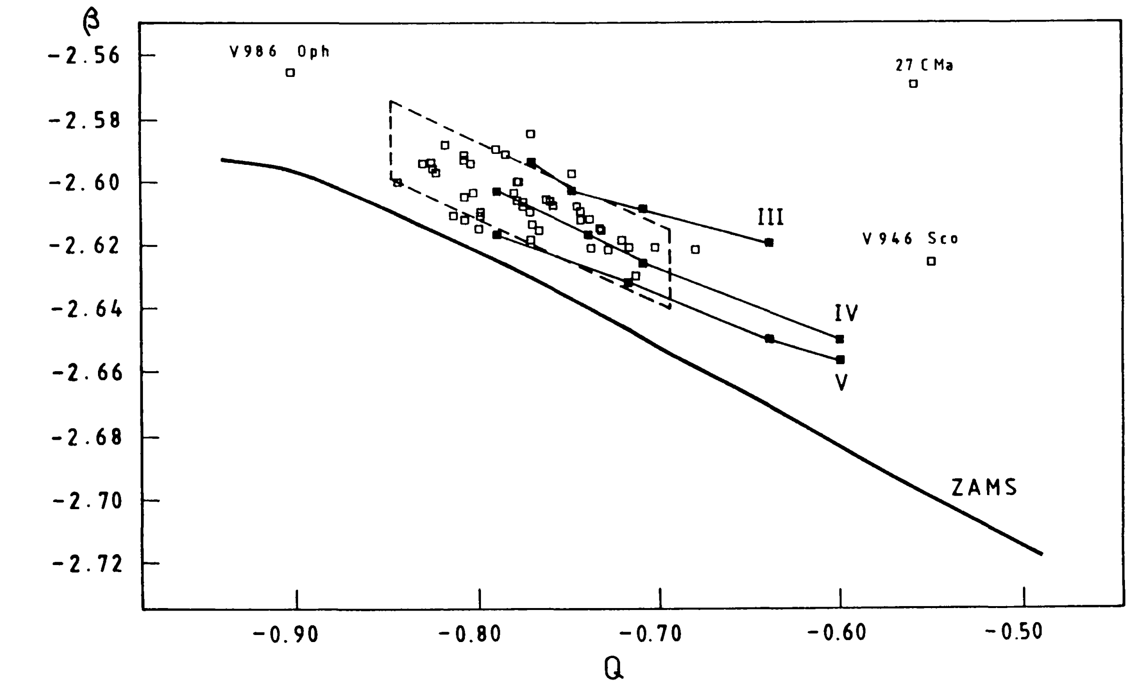

Fig. 6. Updated version of Figure 3. Plotted are all β Cephei stars (open squares) of Table I for which UBV and β photometry is available. Also shown are – transformed from the $c_0 - \beta$ plane – the ZAMS of Crawford (1978) and borders of the β Cephei instability strip (dashed lines), defined by Jerzykiewicz and Sterken (1979) by means of field β Cephei stars (see Figure 8 below).

diagram, occupied by the several longest-known members of the group. Therefore, β Cephei stars were found only in the spectral type range from B0.5 to B2 simply because no one has looked for them elsewhere. In other words, Sareyan *et al.* (1980) believe that the concentration of β Cephei stars in this range has arisen as a result of observational selection.

This interesting suggestion is, however, unfounded. As can be seen from Figs. 4 and 5, Jerzykiewicz and Sterken (1977) included B0 and B2.5 stars in their search program. In addition, they used many B stars from outside the "β Cephei box" as comparison stars. Also Balona and his co-workers (Balona and Engelbrecht 1983, Balona 1983, and Balona and Engelbrecht 1985a) did not limit their observations of the NGC 3293 and NGC 6231 stars to spectral types from B0.5 to B2.

V986 Oph, 27 (EW) CMa, and some of the β Cephei suspects of Table II that fall outside this spectral type range will be discussed in Section 3.7.

3.2. The Q-β diagram

In order to plot β Cephei stars in the observational H-R diagram, Lesh and Aizenman (1973a) used the reddening-free photometric parameters Q and β, which for B stars are related to effective temperature and absolute magnitude, respectively. Lesh and Aizenman's (1973a) diagram was reproduced already in Figure 3.

An updated version of this diagram is presented in Figure 6. The abscissae were computed from the definition

$$Q = (U - B) - \frac{E(U - B)}{E(B - V)}(B - V) \quad (1)$$

where the slope of the reddening line is given by the following expression:

$$\frac{E(U - B)}{E(B - V)} = 0.58 - 0.33\,(B - V)_0 + 0.06\,E(B - V) \quad (2)$$

due to Serkowski (1963). In order to derive $(B - V)_0$ and $E(B - V)$ for use in this expression, we assumed Serkowski's (1963) main-sequence intrinsic two-color relation. We made an attempt to use all available UBV data. The main sources of data included Johnson *et al.* (1966), Crawford *et al.* (1971b), and Deutschman *et al.* (1976) for the field β Cephei stars, Feinstein and Marraco (1980) and Turner et al. (1980) for NGC 3293, Perry *et al.* (1976) for NGC 4755, and Feinstein and Ferrer (1968), Schild *et al.* (1969), and Garrison and Schild (1979) for NGC 6231. If two or more values of the color indices were available for the same star, straight means were used in deriving Q. The UBV color indices and the Q values of the β Cephei stars are given in Table IA of the Appendix.

The β indices were taken mainly from Shaw (1975), Shobbrook (1978a, 1979a), and Olsen (1979) for the field stars, and from Shobbrook (1980, 1983b) for stars in NGC 3293, Perry *et al.* (1976) and Shobbrook (1984) for stars in NGC 4755, and from Crawford *et al.* (1971a) and Shobbrook (1983c) for the NGC 6231 stars. In cases when several values of the β index were available, straight means were used as ordinates in Figure 6. These β indices are listed in Table IIA of the Appendix.

Also shown in Figure 6 are, transformed from the $c_0 - \beta$ plane, the zero-age main-sequence (ZAMS) of Crawford (1978) and borders of the β Cephei instability strip, defined by Jerzykiewicz and Sterken (1979) by means of field β Cephei stars (see Figure 8 below).

In spite of over threefold increase in the number of β Cephei stars since the review paper of Lesh and Aizenman (1978) was written, almost all new β Cephei stars lie within an interval of Q only slightly wider than that occupied by the "confirmed" β Cephei stars in Figure 3.

The stars of Table I which lack $U - B$, $B - V$ and/or β indices, and are therefore not plotted in Figure 6, include KK Vel, IL Vel, V856 Cen, BU Cir, V836 Cen, V349 Nor, V348 Nor, HD 166540 and PHL 346. For KK Vel, V856 Cen, V836 Cen and HD 166540, UBV color indices are available. These four stars have Q values from -0.843 for HD 166540 to -0.768 for V856 Cen, that is, falling within the interval of Q indicated in Figure 6 by vertical dashed lines. In particular, V836 Cen, mentioned in the preceding section as one of the B3 stars in Figure 5, has $Q = -0.748$. This value corresponds to spectral type close to B1.5[2].

[2] The UBV photometry and the spectral classification of V836 Cen are both due to Hill (1970). He finds that a mean difference between his spectral classification and that of other workers amounts to +0.5 of a sub-type, with a standard deviation for one star equal to ± 1.1. The difference between Hill's (1970) B3 and the photometric B1.5 is not grossly inconsistent with these numbers.

Thus, almost all β Cephei stars for which UBV indices are available have Q falling within the interval $-0.85 < Q < -0.70$. This is, of course, another version of the conclusion reached in the preceding section, somewhat more precise, because expressed in terms of a continuous parameter, instead of the "quantized" spectral type.

Three β Cephei stars appear to lie far off the limits of the above-mentioned interval of Q. One of them, V946 Sco = NGC 6231-261, has a Q value inconsistent with its MK type of B2 IV-Vn. However, this MK type agrees very well with the star's X and c_0 (see below, Sections 3.4 and 3.5). Apparently, UBV color indices of V946 Sco, based on a single observation (Feinstein and Ferrer 1968), are in error.

The other two stars, V986 Oph and 27 (EW) CMa, cannot be dismissed so easily. The UBV and Strömgren color indices of V986 Oph are well determined and its Q and c_0 (see the next section) agree with each other, although both are slightly too small for the star's MK type of B0 IIIn. In the case of 27 (EW) CMa, the Q and X (see Section 3.4) are consistent with the B3 spectral type, but c_0 indicates a spectral type of about B2 (see the next section). The β index is clearly affected by emission. We shall return to V986 Oph and 27 (EW) CMa in Section 3.7.

Except for 27 (EW) CMa, the vertical spread of the β Cephei stars in Figure 6 is only slightly greater than that in Figure 3. The width of the β Cephei strip in β will be discussed in the next section.

3.3. THE STRÖMGREN $uvby\beta$ SYSTEM

As can be seen from eq. (1), Q is dominated by $U - B$. Hence, it suffers from the well-known difficulties of correcting the broad-band ultraviolet observations for atmospheric extinction and transforming them to the standard system (see, e.g., Hardie 1966). These difficulties are much reduced in the case of intermediate-band ultraviolet color-indices such as $u - b$, although they are not always insignificant (see Section 3.5). In addition, the intermediate-band ultraviolet indices are more sensitive than Q to the size of the Balmer jump.

A number of workers investigated the position of β Cephei stars in the observational H-R diagram, selecting as the abscissa one of the de-reddened color-indices involving the u band, that is, $(u - b)_0$, c_0 , or $[u - b]$. Shobbrook (1976) and Davis and Shobbrook (1977) give reasons for preferring c_0 to $[u - b]$ as the temperature parameter.

For an early-type star, c_0 is derived from the observed $c_1 = (u - v) - (v - b)$ by means of an iterative procedure involving the intrinsic two-color relation, $(b-y)_0 = 0.099\, c_0 - 0.117$, and the slope of the reddening line, $E(c_1)/E(b - y)$, assumed to be equal to 0.24 (Crawford 1973, Shobbrook 1978a). Unless stated otherwise, all c_0 values quoted in the present paper were obtained in this way.

Shaw (1975) carried out $uvby\beta$ photometric observations of 9 β Cephei stars and 22 β Cephei suspects. From these observations, supplemented by data taken from the literature, he constructed an observational H-R diagram using $(u - b)_0$

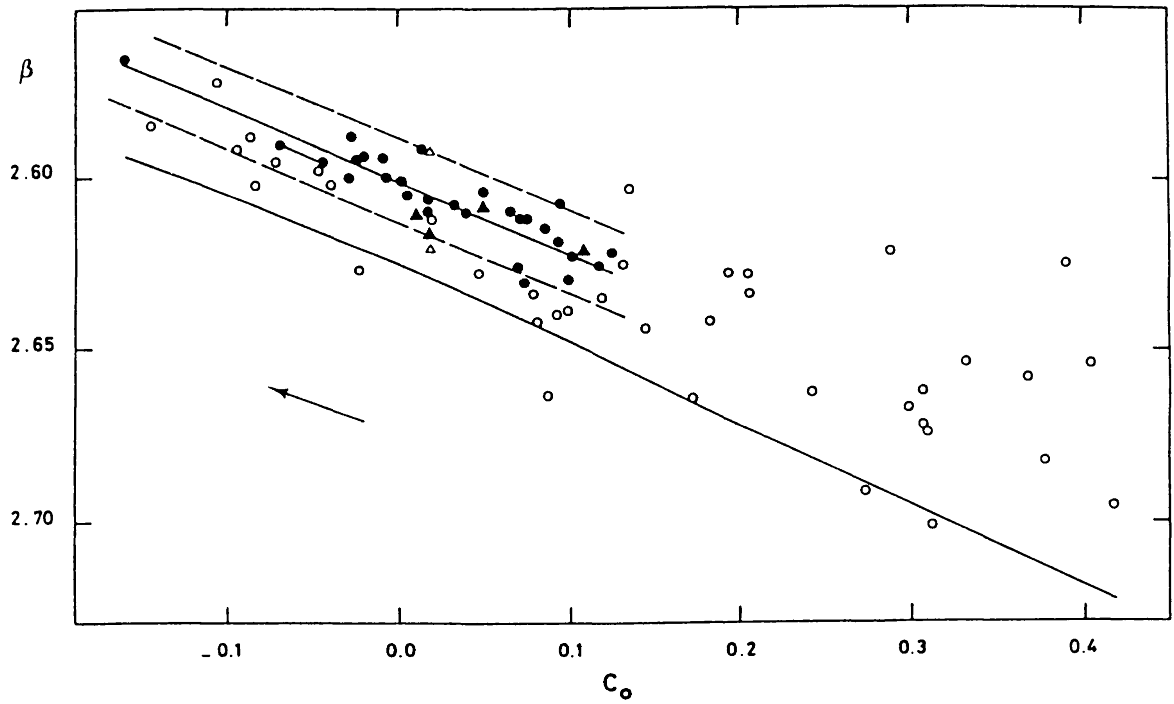

Fig. 7. The field β Cephei variables (filled circles and triangles) and nonvariable early B stars (open circles and triangles) in the $c_0 - \beta$ plane (from Jerzykiewicz and Sterken 1979). The four β Cephei stars, discovered by Jerzykiewicz and Sterken (1977, 1979), are shown as filled triangles; the three leftmost ones are the shortest-period field variables mentioned in the text. Open triangles correspond to the two available values of β for HR 3453. The arrow indicates a maximum duplicity correction as estimated by Shobbrook (1978a). The two filled circles at the left and right ends of the short straight line-segment correspond to β Cru, uncorrected and corrected for duplicity, respectively. The lower solid line is Crawford's (1978) ZAMS, and the upper one defines the ridge line of the β Cephei instability strip. The dashed lines indicate the lower and upper boundaries of the strip.

and β as coordinates. He noted that almost all β Cephei stars fall into the $(u-b)_0$ range from approximately -0.2 to 0.1, and that they occupy a strip with limited width in β, similar to that defined earlier by Lesh and Aizenman (see Figure 3). A notable exception was τ^1 Lup, for which Shaw quotes $\beta = 2.739$ from Crawford *et al.* (1970). This value would place the star about one-tenth of a magnitude below the ZAMS. However, Grønbech and Olsen (1976) give β = 2.625, which brings the position of τ^1 Lup in accord with other β Cephei stars. Subsequent extensive observations of Shobbrook (1978a) did not, unfortunately, resolve this discrepancy. Shobbrook (1978a) summarizes his results as follows:

> "During 1975 and 1976, the value most often observed was close to 2.620, but over two nights the value averaged 2.745, which is more appropriate for a late B star. Intermediate values were not observed."

On another occasion, the same observer found a mean β = 2.622. He concluded that the higher value must occur only rarely (Shobbrook 1979a). In the present paper we disregard the higher value altogether.

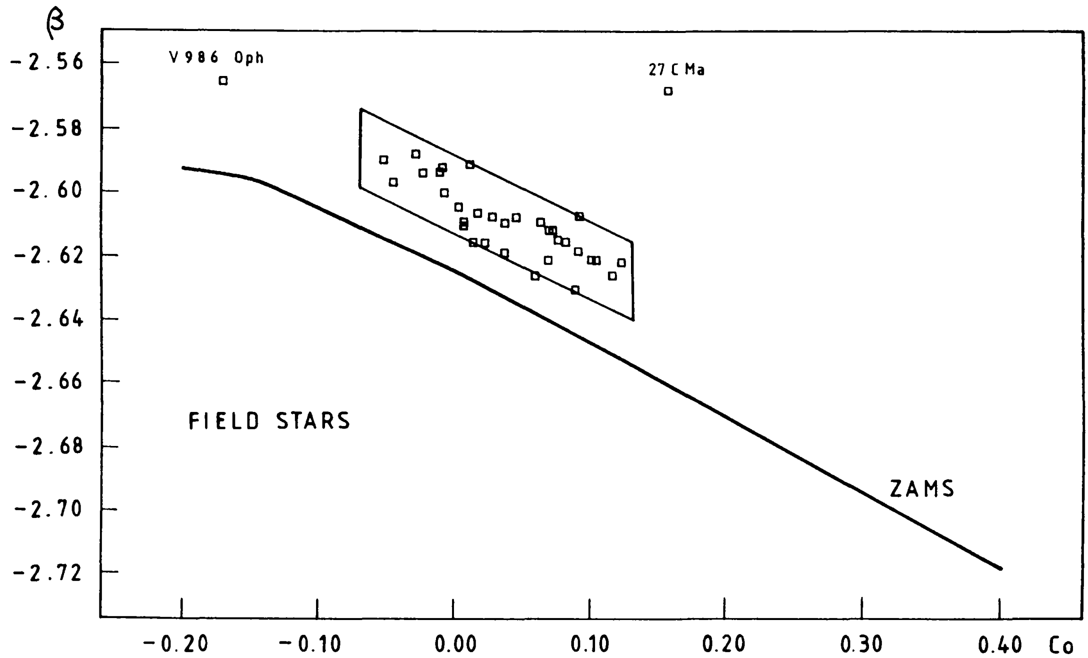

Fig. 8. The field β Cephei variables of Table I in the $c_0 - \beta$ plane. No duplicity corrections were taken into account. ZAMS and the β boundaries of the instability strip are the same as in Figure 7. The limits of the strip at $c_0 = -0.07$ and $c_0 = 0.13$ are also indicated.

The behaviour of τ^1 Lup is unique. Most other β Cephei variables are constant in β to better than $0^m_.010$. This conclusion is due to Shobbrook (1978a), who obtained β values accurate to $\pm 0^m_.002$ or $\pm 0^m_.003$ for all bright early B stars observable from Siding Spring Observatory. Using these data, published β values for stars he could not reach from Siding Spring, and published *uvby* photometry, Shobbrook (1978a) found that in the $c_0 - \beta$ plane the width of the β Cephei instability strip does not exceed $0^m_.003$ in β. He concluded that the strip is so narrow that it is not even resolved by the observations. In addition, Shobbrook (1978a) found several nonvariable stars in the instability strip. He estimated (Shobbrook 1978b) that about three-quarters of the stars in the strip, to a distance modulus of $8^m_.0$, are β Cephei variables.

As far as the width of the instability strip is concerned, Shobbrook's (1978a) findings were contradicted by Jerzykiewicz and Sterken (1979) and Sterken and Jerzykiewicz (1980). Like Shobbrook (1978a), they investigated the position of β Cephei variables and nonvariable early B stars in the $c_0 - \beta$ plane. They used all available β values, including those of Shobbrook (see Section 3.2 above for other references), and *uvby* photometry from Olsen (1979). The result of Jerzykiewicz and Sterken (1979) is reproduced here in Figure 7. As can be seen from this figure, the instability strip has a substantial width along β, amounting to $0^m_.025$, almost ten times that found by Shobbrook (1978a).

The main reason for this discrepancy comes probably from the fact that Shobbrook could not take into account β Cephei stars which Jerzykiewicz and Sterken (1979) used for determining the β boundaries of the instability strip, because their variability was discovered after Shobbrook's analysis had been completed. For example, 19 (V637) Mon, lying in Figure 7 on the upper boundary at $c_0 = 0.012$, and HR 3058 = QS Pup, lying on the lower boundary at $c_0 = 0.014$, were both regarded by Shobbrook (1978a) to be nonvariable. The latter star, together with HR 2928 = PT Pup and HR 3088 = V372 Car (the three leftmost filled triangles in Figure 7), are the shortest-period field β Cephei stars, mentioned in Section 2.1. Note, however, that even the "confirmed" β Cephei stars in Figure 3 show significant spread in β. Subsequent observations of β Cephei stars in open clusters (see Section 3.5) provided conclusive evidence against the notion of the infinitesimal width of the β Cephei instability strip.

All β Cephei stars in Figure 7, except V986 Oph (filled circle at the extreme upper left), fall within the interval $-0.07 < c_0 < 0.13$. In this respect, Jerzykiewicz and Sterken (1979) agree with Shobbrook (1978a). However, they find only one nonvariable star in the above-mentioned interval in their wide instability strip, whereas Shobbrook (1978a) finds several in his narrow one. We shall return to the question of nonvariable stars in the instability strip in Section 3.6.

An updated version of Figure 7 is presented in Figure 8. It shows all field β Cephei stars of Table I except KK Vel, V856 Cen, V836 Cen, V349 Nor, V348 Nor, V831 Ara, HD 166540 and PHL 346, which lack $uvby\beta$ photometry. The $uvby$ data used to derive c_0 indices of stars plotted in Figure 8 were taken mainly from Olsen (1979) and Shaw (1975). These c_0 values are listed in Table IIA of the Appendix. The lower and upper boundaries of the instability strip in Figure 8 are the same as in Figure 7. In addition, the above-mentioned c_0 limits at -0.07 and 0.13 are indicated. As we already mentioned in the preceding section, 27 (EW) CMa has a c_0 value indicating a B2 spectral type.

Another study of the position of field β Cephei stars in the color-magnitude diagram was undertaken by Jakate (1979a). He used $[u-b] = (u-b) - 1.84\,(b-y)$, Crawford's analogue of Q, as the effective-temperature parameter. Including HR 3924 and HR 3941 = IV Vel, two stars which he found to be short-period variables (see Table II), Jakate (1979a) concluded that the instability strip defined by Shobbrook (1978a) should be extended to cooler temperatures and to lower luminosities. A low-temperature extension of the β Cephei instability strip has also been advocated by Sareyan et al. (1979) and, more recently, by Chapellier *et al.* (1987). We shall return to this point in Section 3.7.

Strömgren photometry was also extensively used to study the position in the H-R diagram of β Cephei stars which are members of galactic clusters. This topic will be treated in Section 3.5.

3.4. The Geneva photometric system

Using linear combinations of the Geneva photometric system color-indices, Cramer and Maeder (1979) defined a set of reddening-free orthogonal coordinates in such a way that its X axis was oriented along the main sequence of the O and B stars, Y pointed in the direction of increasing luminosity, and Z was significantly different

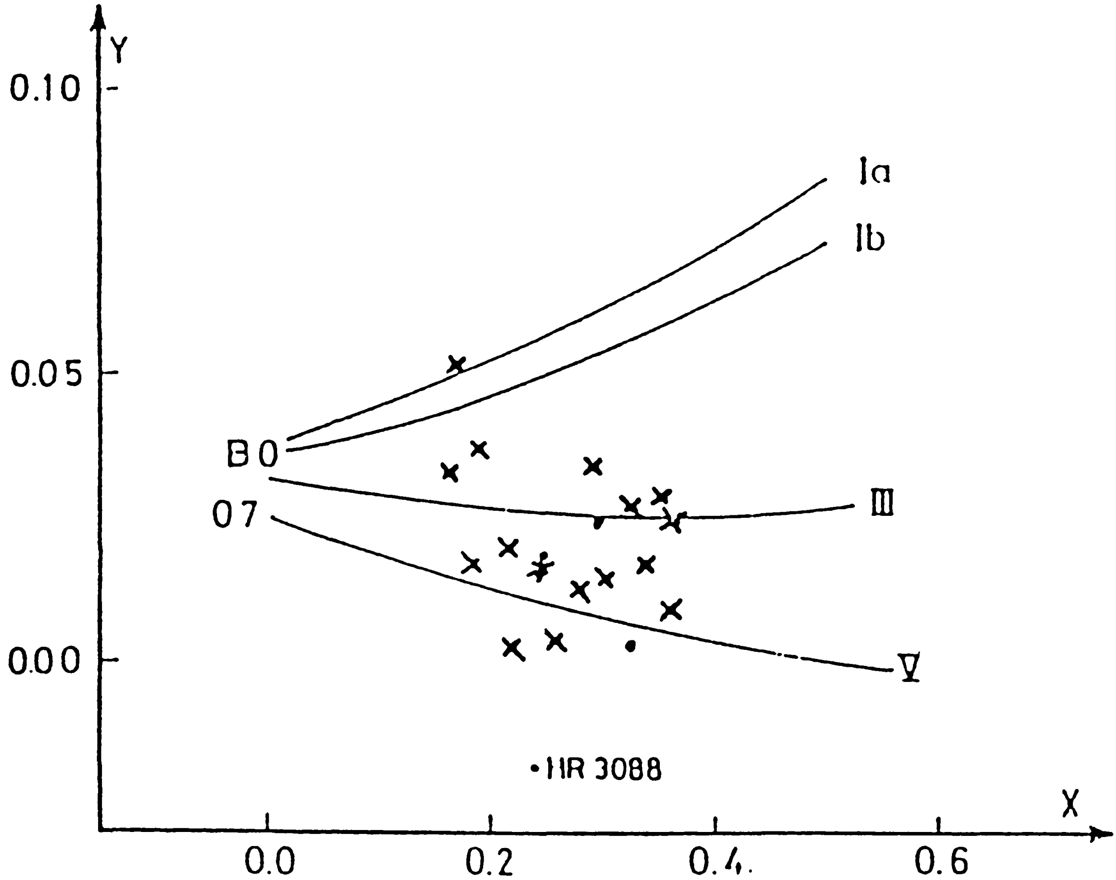

Fig. 9. The confirmed β Cephei stars (crosses) and β Cephei candidates (dots) of Lesh and Aizenman (1978) in the Geneva system $X - Y$ plane, and average sequences for MK luminosity classes Ia, Ib, III and V (from Waelkens 1981).

from zero only for chemically peculiar B and A stars. For the B stars, the $X - Y$ diagram is, in principle, equivalent to the $c_0 - \beta$ diagram. However, Y has an important advantage over β as an absolute-magnitude parameter, because it is relatively insensitive to hydrogen-line emission. This follows from the fact that as a measure of absolute magnitude Y uses the intensity of high-order Balmer lines, which are much less affected by emission than Hβ.

A first attempt to investigate the position of β Cephei stars in the $X - Y$ plane was undertaken by Waelkens (1981). In his $X - Y$ diagram, reproduced in Figure 9, Waelkens (1981) plotted the confirmed β Cephei stars and β Cephei candidates of Lesh and Aizenman (1978). He used the color indices from the third edition of the Geneva catalogue by Rufener (1980). Subsequently, Waelkens (1984), Waelkens and Cuypers (1985), Waelkens and Rufener (1988), Waelkens and Heyndericks (1989), and Waelkens *et al.* (1991) obtained X and Y coordinates for a number of β Cephei stars not shown in Figure 9. The color indices on which these data have been based were superseded by those in the fourth edition of the Geneva catalogue (Rufener 1988). Altogether, 51 stars of Table I appear in this catalogue. An $X - Y$ diagram showing these stars is presented in Figure 10, and the X, Y, Z values are

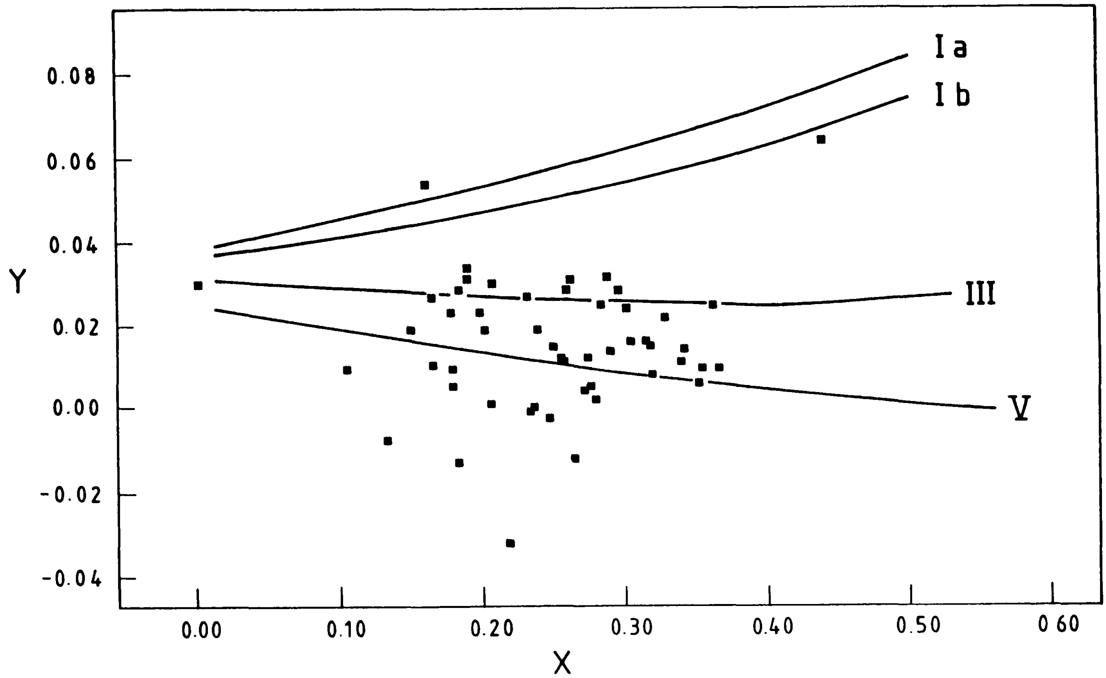

Fig. 10. The same as Figure 9 for all β Cephei stars of Table I that appear in the fourth edition of the Geneva catalogue (Rufener 1988).

given in Table IIIA of the Appendix.

Except for V986 Oph (the leftmost square), all β Cephei stars in Figure 10 fall within the interval $0.10 < X < 0.37$. By means of a correlation between c_0 and X, easily obtained for B stars measured in both systems, this interval of X can be translated into $-0.09 < c_0 < 0.12$. In comparison with the result of Jerzykiewicz and Sterken (1979), mentioned in the preceding section, both sides of this inequality are slightly bluer, but there is no serious disagreement. This is important, because Figure 10 includes V349 Nor, V348 Nor, and PHL 346, which lack both UBV and Strömgren photometry, and were therefore not represented in Figs. 6 and 8.

V836 Cen, the B3 V star in Figure 5 (see Section 3.1), has $X = 0.272$ and $Y = 0.004$. These values indicate an MK type close to B1 V (see Cramer 1984, Table IVa), in good agreement with the spectral type corresponding to Q (see Section 3.2). On the other hand, the X coordinate of 27 (EW) CMa (the rightmost square) approximately agrees with an average value for the MK type of B3 III (see Cramer 1984, Table IVb). However, the star's Y coordinate is discordant (see below).

The square at $X = 0.265$ and $Y = -0.012$, below the luminosity class V average sequence, is V946 Sco, the star with the discordant Q value (see Figure 6). The X coordinate is consistent with the star's spectral type of B2, but the Y coordinate does not agree with its luminosity class of IV-V.

A number of β Cephei stars in Figs. 9 and 10 lie outside the average sequences

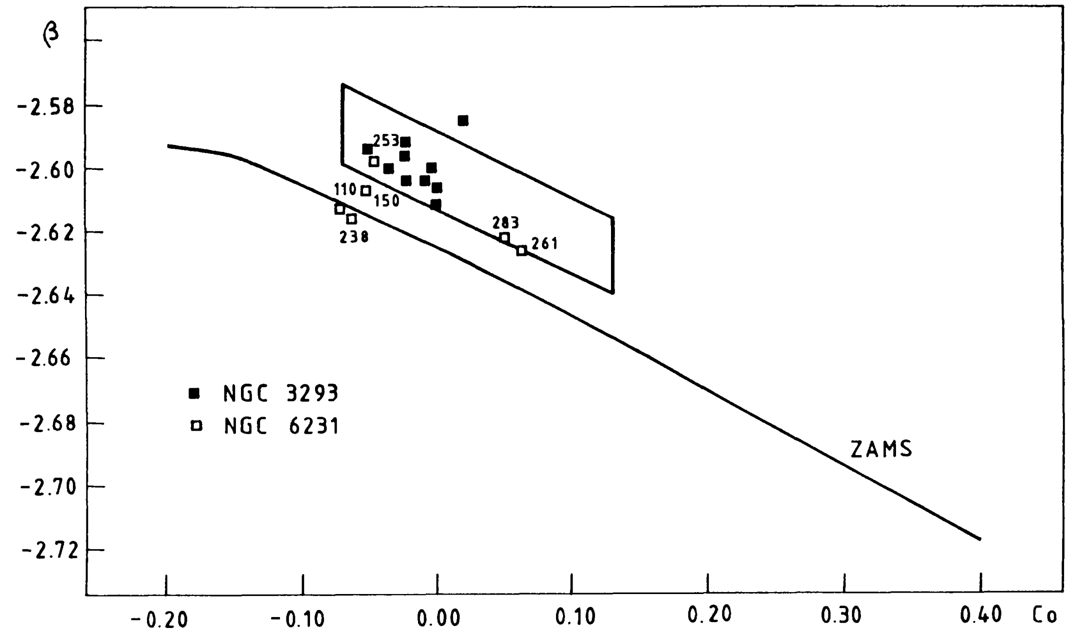

Fig. 11. NGC 3293 and NGC 6231 β Cephei stars in the $c_0 - \beta$ plane. No duplicity corrections were taken into account. ZAMS and borders of the instability strip are the same as in Figure 8.

for luminosity classes V and III. In this respect the $X - Y$ diagram differs strikingly from the pattern seen in Figure 6. A part of this difference may be due to errors of Y, perhaps caused by inadequate averaging over the cycle of short-period variability of these stars. Still, the Y coordinates of σ Sco and 27 (EW) CMa (the square above the Ia average sequence, and the one just below the Ib sequence, respectively) require accounting for. In the case of σ Sco, Waelkens and Cuypers (1985) suspect influence of circumstellar material on the star's Geneva color indices. The same effect may be responsible for the Y coordinate of 27 (EW) CMa. To what extent the Y coordinates of other stars in Figure 10 are affected by residual interstellar extinction effects is not known. In any case, the usefulness of Y as a measure of absolute magnitude of β Cephei stars appears to be somewhat limited at present. This is rather unfortunate in view of the small sensitivity of Y to hydrogen-line emission, mentioned in the opening paragraph of this section.

3.5. The Evolutionary State of β Cephei Stars in Open Clusters

Figure 11 shows the NGC 3293 and NGC 6231 β Cephei stars in the $c_0 - \beta$ plane. BW Cru, the only confirmed β Cephei variable in NGC 4755, is not plotted because its position coincides with that of V406 Car = NGC 3293-18 at $c_0 = 0.000$ and $\beta =$ 2.606. The filled square above the upper boundary of the instability strip represents V412 Car = NGC 3293-65, sometimes also referred to as NGC 3293-C. The β index used to plot the star in Figure 11 is a mean of two discordant values, both obtained by Shobbrook (1980, 1983b). The rightmost open circle represents V946

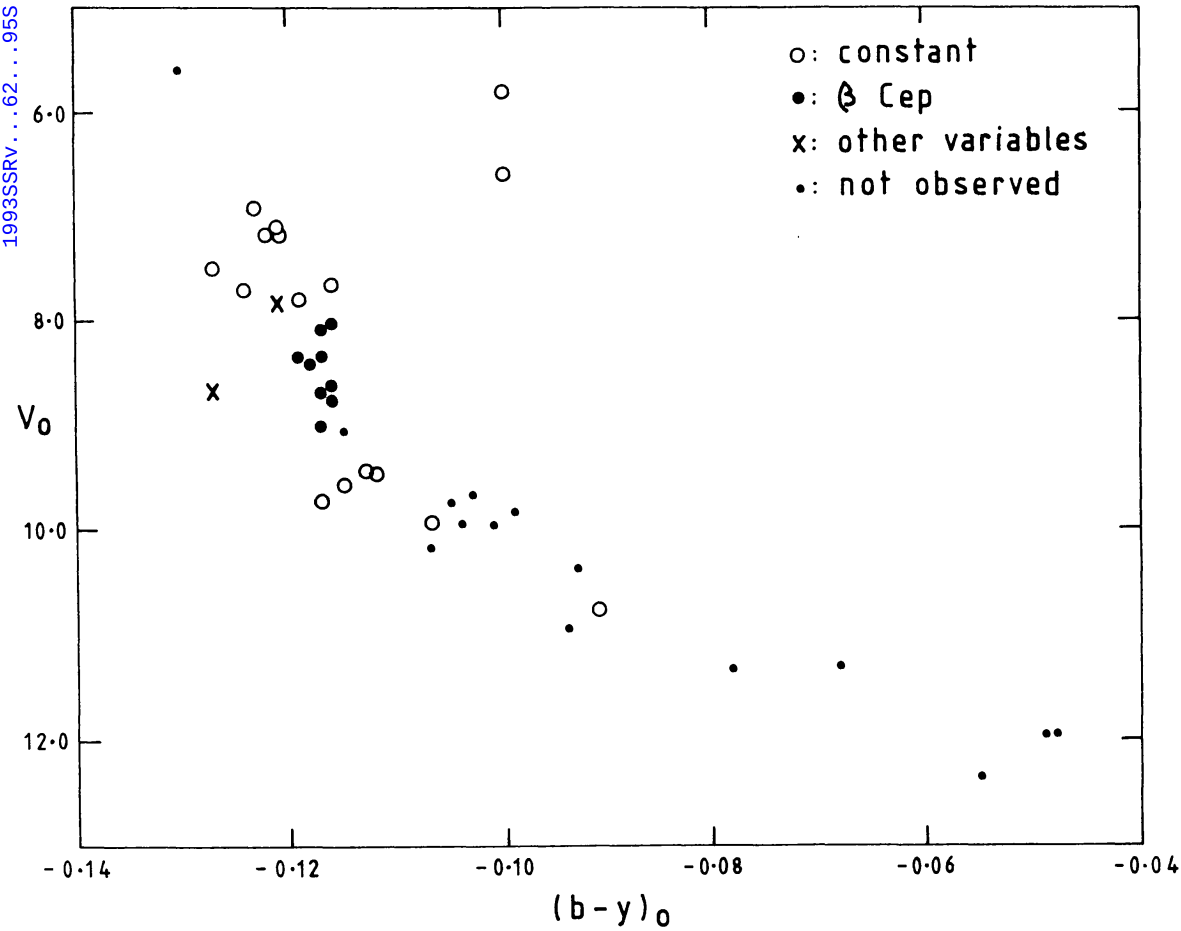

Fig. 12. The color-magnitude diagram of NGC 3293 showing the sequence of β Cephei stars (from Balona and Engelbrecht 1981).

Sco = NGC 6231-261, the star with a discordant Q value (see Figure 6). The star's position in Figure 11 is consistent with its MK type of B2 IV-V.

In the $c_0 - \beta$ plane, the NGC 3293 β Cephei stars, except V412 Car, form a sequence over limited range of c_0 within the instability strip, defined by field β Cephei variables. The sequence can also be seen in the color-magnitude diagram of NGC 3293, reproduced here – in Figure 12 – from an early paper by Balona and Engelbrecht (1981). In this diagram, based on Shobbrook's (1980) detailed investigation of the reddening and distance modulus of the cluster, V412 Car is plotted as a "not observed" star at the red end of the β Cephei sequence. Its β Cephei-type variability, suspected by Balona and Engelbrecht (1983), was confirmed by Engelbrecht (1986). The star seems to have a peculiar β index.

As can be seen from Figure 12, all NGC 3293 β Cephei stars lie on the cluster's evolved main-sequence. This important discovery led Balona and Engelbrecht (1981) to the conclusion that these variables are in the late stages of core hydrogen-burning, just preceding the secondary gravitational contraction. As Balona and Engelbrecht (1981) point out, their conclusion is independent of the uncertain process of matching theoretical evolutionary tracks with observations.

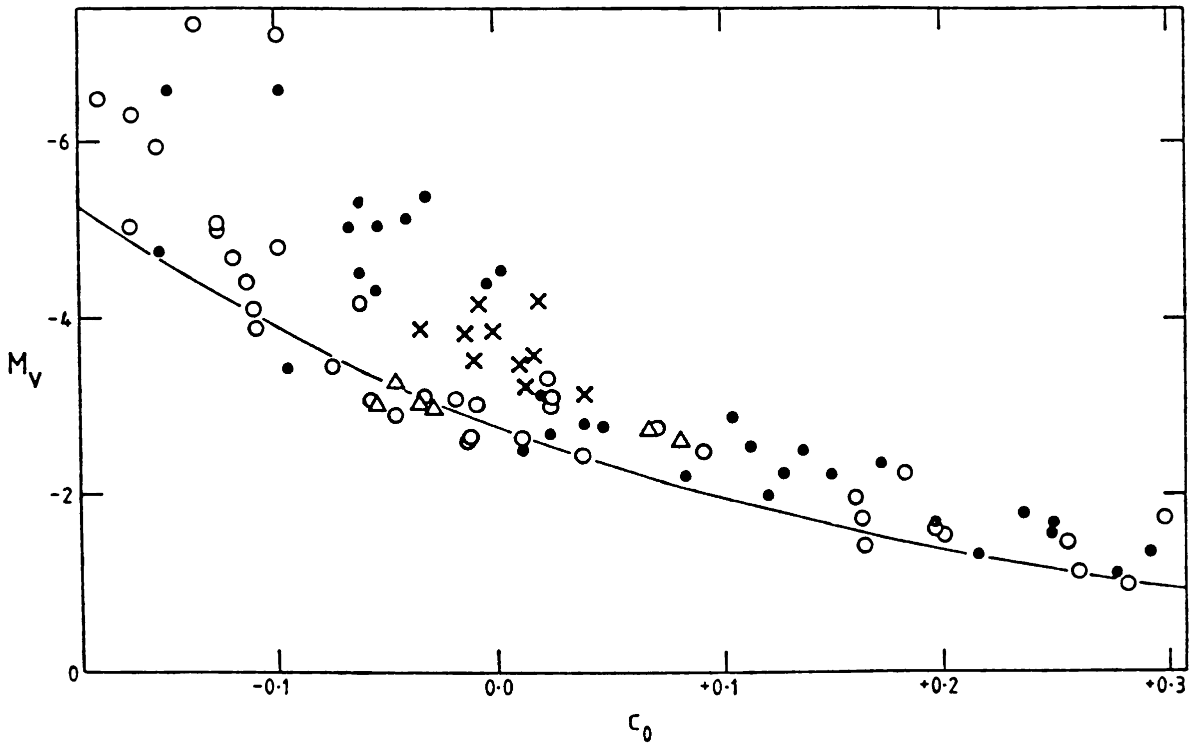

Fig. 13. Observational H-R diagram for NGC 3293 and NGC 6231 (from Balona and Engelbrecht 1985a). Crosses and triangles represent β Cephei stars in NGC 3293 and NGC 6231, respectively. Other stars are plotted as filled circles (NGC 3293), and as open circles (NGC 6231). The ZAMS, shown as a solid line, runs slightly below that of Crawford (1978) for $-0.07 < c_0 < 0.15$; the difference amounts to about $0\overset{\mathrm{m}}{.}2$ in the middle of this interval.

Shobbrook's (1979a) discovery of HDE 326333 in the young cluster NGC 6231, and the subsequent discovery of five other β Cephei variables in this cluster by Balona (1983) and Balona and Engelbrecht (1985a), have shown that the β Cephei phenomenon occurs also close to the ZAMS. This can be seen from the c_0 – M_V diagram in Figure 13, reproduced from Balona and Engelbrecht (1985a), where the NGC 6231 and NGC 3293 stars are plotted together using a distance modulus of $11\overset{\mathrm{m}}{.}21$ for NGC 6231, and 12.15 for NGC 3293 (Balona and Shobbrook 1984). As Balona and Engelbrecht (1985a) put it, "the substantial difference in evolutionary age between β Cephei stars in the two clusters is evident." Note, however, that an answer to the question how large this difference actually is depends on uncertainties of the main-sequence fitting procedure. Unfortunately, in the c_0 – β diagram (Figure 11), which is, of course, free from these uncertainties, the four hottest NGC 6231 β Cephei variables show an uncomfortably large spread in β.

Despite these difficulties, the discovery of β Cephei stars in NGC 6231 led Balona and his co-workers to the conclusion that – to use Balona's (1987a) words – "β Cephei variability is probably found from the ZAMS until the end of core hydrogen-burning."

As far as the exact limits of the c_0 interval occupied by β Cephei stars are concerned, one should realize that unexpectedly large systematic effects may occur because of improper selection of standard stars or because of inconsistent filter passbands. This was demonstrated by Sterken and Manfroid (1987), who determined Strömgren color indices of β Cephei stars in NGC 3293 using different *uvby* filter sets. From a comparison of their results with those reported earlier by Shobbrook (1980, 1983b), Sterken and Manfroid (1987) concluded that the position in the $c_0 - \beta$ plane of β Cephei stars, belonging to a reddened cluster such as NGC 3293, may be strongly biased by systematic errors in c_0.

3.6. Nonvariable Stars in the Instability Strip

Early attempts to answer the question whether β Cephei variables and nonvariable stars coexist in the instability strip, based on observations of field stars, led to conflicting results. While Watson (1972) – who was probably the first to address the question – believed that β Cephei variables and constant stars separate in the H-R diagram, Lesh and Aizenman (1973a) found many nonvariable B stars in the region of the $Q - \beta$ plane occupied by β Cephei variables. Watson's (1972) work was criticized by Lesh and Aizenman (1973b) on the grounds that only four of his 28 nonvariable stars fell in the same spectral-type range as the β Cephei variables. However, the conclusion of Lesh and Aizenman (1973a) is also questionable because most stars they considered constant were not checked for variability.

In order to improve the situation, Jerzykiewicz and Sterken (1979) compiled a list of 40 early-type stars, found constant in light by Lynds (1959), Jerzykiewicz (1992), and Jerzykiewicz and Sterken (1977). These stars have been plotted in Figure 7. From this figure, Jerzykiewicz and Sterken (1979) concluded that constant stars probably separate from β Cephei variables in the interval $-0.07 < c_0 < 0.13$. Note, however, that some nonvariable stars lie in Figure 7 below the lower boundary of the instability strip, defined by the field variables. Later, Sterken and Jerzykiewicz (1983) found more nonvariable field stars in the interval $-0.07 < c_0 < 0.13$, a few also above the lower boundary of the instability strip (see Figure 14 in the next section). These facts, together with the discovery of β Cephei stars in NGC 6231, which requires shifting the lower boundary of the instability strip toward the ZAMS (see the preceding section), contradict the above-mentioned conclusion of Jerzykiewicz and Sterken (1979).

In NGC 3293, constant stars are found only outside the β Cephei sequence in the color-magnitude diagram (Figure 12). Apparently, all stars which fall into the instability strip pulsate. The hottest β Cephei star in this cluster, V378 Car, has $c_0 = -0.051$, a value close to c_0 indices of the hottest field variables with the exception of V986 Oph. On the other hand, the cool end of the sequence at $c_0 = 0.02$ is much hotter than the low-temperature limit of the instability strip in Figure 11. It would thus seem that NGC 3293 stars cooler than $c_0 = 0.02$ are constant because they have not yet reached the instability strip. This conclusion, due to Balona and Engelbrecht (1983), has been revised by the same authors (Balona

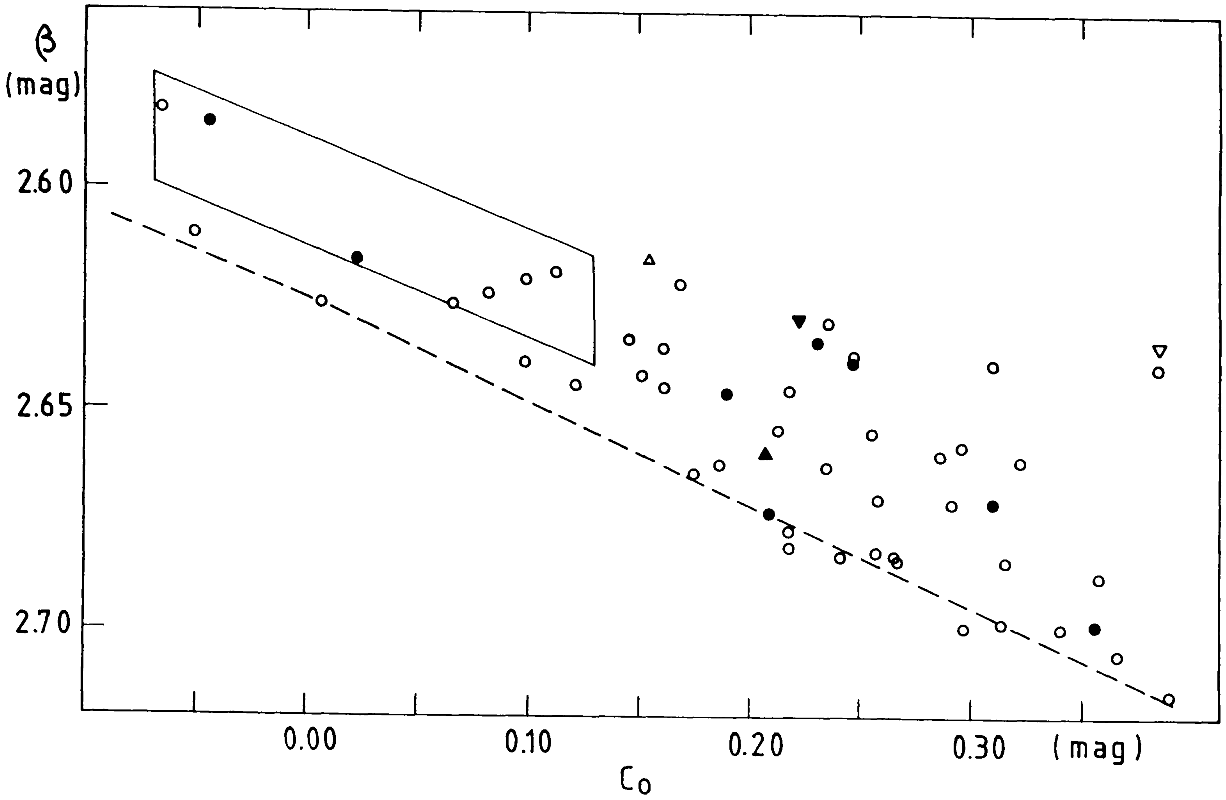

Fig. 14. A sample of bright early B stars, found variable and nonvariable in light (filled and open symbols, respectively), shown in the c_0 - β plane (from Sterken and Jerzykiewicz 1983). The ZAMS (dashed line) and borders of the instability strip (solid lines) are the same as in Figure 8.

and Engelbrecht 1985a) following Balona's (1983) discovery that in NGC 6231 nonvariable stars occur in the same interval of c_0 as β Cephei variables. Balona and Engelbrecht (1985a) conclude that β Cephei and non-β Cephei stars can coexist in the instability strip, and that about half of the B0 to B2 dwarfs and giants are likely to be β Cephei variables. From the above discussion it follows also that the incidence of nonvariable stars in the instability strip increases toward the ZAMS.

Interpretation of these purely observational conclusions is complicated because the concept of nonvariable star is so vague. In the first place, there are differences in photometric precision. For example, Balona and Engelbrecht (1985a) set an upper limit of $0^{\mathrm{m}}.0015$ on semi-amplitude for stars they found constant in NGC 6231, whereas Sterken and Jerzykiewicz (1983) estimate the detection threshold of their search program to be equal to $0^{\mathrm{m}}.0050$.

In addition, the large number of very-small-amplitude variables among β Cephei stars suggests that there may be stars in the instability strip with undetectably small light variations. It is not known how many stars belong to this category. These apparently constant stars may include radial and low l nonradial pulsators having genuinely small surface amplitudes, as well as high l pulsators, with variations masked by the effect of averaging over the stellar disk.

Another type of nonradial pulsator with undetectable light variations may be represented by the primary component of the double-line spectroscopic binary α Vir (Spica). The star was found by Shobbrook *et al.* (1969) to show β Cephei-type light variation with a period of $0^{d}.1738$ and a yellow light amplitude of $0^{m}.016$, in addition to an ellipsoidal variation with the system's 4-day orbital period and a range of $0^{m}.03$. The $0^{d}.1738$ variation in the star's radial-velocity data, published by various observers since about 1910, has been uncovered by Smak (1970). However, in 1972 the short-period variations become undetectable (see Lomb 1978 for a thorough study of the decline of the star's pulsations). Recent observations by Sterken *et al.* (1986) and Balona (1989) show that only the 4-day ellipsoidal light-variation can now be seen.

Balona (1985) suggested that pulsations of the primary of Spica are now undetectable because of an unfavourable change in the inclination of the pulsation axis, caused by precession in the binary system. According to Balona (1985), the observed decline of Spica's short-period variations can be accounted for by assuming a nonradial quadrupole mode and a period of precession equal to about 200 years.

A small fraction of nonradial pulsators in the instability strip may appear nonvariable because of this effect. While, however, the hypothesis can be verified in the case of a binary system such as Spica, since the variations would periodically reappear as the inclination angle changes because of precession, it is not verifiable in the case of single stars.

Among β Cephei suspects, listed in Table II, there are stars that were found nonvariable or variable on long time-scales, following earlier reports of their short-period variability. In most such cases the follow-up observations were of higher quality and more numerous than the original ones, suggesting that the original evidence for short-period variability has been inadequate. A good example is 53 (UW) Ari. This B1.5 V runaway star has c_0 within the range defined by field β Cephei variables, but its β index places it below the ZAMS. 53 Ari has been classified as a β Cephei star on the strength of radial velocity and photometric evidence for short-period variability, provided by van Hoof and Blaauw (1964) and by Bondal (1967). However, the star was shown to be constant in both radial velocity and light by Sterken (1988a), who based this conclusion on his own observations and on reanalysis of all older data.

3.7. High- and low-temperature extensions of the instability strip

In the observational H-R diagrams considered in this section, V986 Oph and 27 (EW) CMa fall outside the instability strip defined by other β Cephei variables. V986 Oph lies near the high-temperature projection of the ridge line of the instability strip. Its spectral type and the various ultraviolet-blue color indices yield an effective temperature about 5000 K higher than the high-temperature boundary of the instability strip in Figs. 6 and 8. On the other hand, the effective temperature of 27 (EW) CMa may be lower than the low-temperature boundary of the strip by as much as 3500 K, if the spectral type and the Q and X indices of this Be star can be trusted. A question arises whether the instability strip should be extended in effective temperature to embrace these two stars, or whether they show β Cephei pulsations because of circumstances that occur in them, but not in other β Cephei stars.

The second possibility seems more likely. Both stars are certainly peculiar. Both are fast rotators, with $v \sin i$ equal to 434 km s^{-1} and 139 km s^{-1} for V986 Oph and 27 (EW) CMa, respectively (Hoffleit 1982), although only the second star has been observed to display a Be spectrum. In addition to β Cephei-type variations, both show longer-period light variations that may be related to rotation (Cuypers *et al.* 1989, Balona and Rozowsky 1992).

An important aspect of 27 (EW) CMa consists in it being the only known object in which a β Cephei variation was seen *in statu nascendi*. As found by Balona and Rozowsky (1992) from extensive observations with a Strömgren β filter, in 1986 and 1987 no short-period variations were present with an amplitude exceeding 2 millimags, while a $0\overset{d}{.}0918$ variation with an amplitude of 3.8 millimag was found in 1990. In 1991 its amplitude increased to 4.3 millimag. It seems that, as expressed by Balona and Rozowsky (1992), "we have witnessed for the first time the birth of β Cephei pulsations."

A number of stars in Table II falls to the right of the low-temperature boundary of the instability strip. Three examples are shown in Figure 14, viz., HR 2603 = HH CMa (open triangle), HR 3924 (filled inverted triangle), and HR 3941 = IV Vel (open inverted triangle). However, according to Sterken and Jerzykiewicz (1983), HH CMa and IV Vel are constant, whereas HR 3924 is variable on a time scale of days.

As we mentioned in Section 3.3, the discovery of short-period light variations of IV Vel and HR 3924 led Jakate (1979a) to the conclusion that the β Cephei instability strip should be extended to cooler temperatures and to lower luminosities. A low-temperature extension of the β Cephei instability strip has also been advocated by Sareyan *et al.* (1979) and, more recently, by Chapellier *et al.* (1987). Sareyan et al. (1979) postulated the extension in order to account for short-period variations of 53 (AG) Psc, while Chapellier *et al.* (1987) added ι Her as another example of a low-temperature β Cephei star. Both stars have c_0 close to 0.3.

For none of these stars were the short-period variations shown to persist over more than one cycle. On the other hand, the above-mentioned negative evidence in the case of IV Vel and HR 3924 has been based on several closely-spaced nights, making an explanation of nonvariability in terms of destructive interference of harmonic components rather unlikely. The same argument applies to 53 Psc, for which there is evidence for long periods of constancy (Jerzykiewicz and Sterken 1989). Perhaps the reported variations are of an "on-and-off" type.

4. β Cephei stars in the theoretical H-R diagram

Before β Cephei stars had been found in NGC 3293 and NGC 6231, there were numerous attempts to determine the evolutionary state of field β Cephei stars. Two methods were used. The first method relied on matching the position of the variables in the theoretical H-R diagram with evolutionary tracks of models of massive stars. The second method involved comparing observed secular period-changes with predictions of the theory. In this section we review results obtained by both these methods. First, however, we shall discuss effective temperatures, bolometric corrections, and absolute magnitudes of stars of early B spectral type.

4.1. Effective Temperatures and Bolometric Corrections

A few β Cephei variables are among the 32 stars for which Code et al. (1976) determined empirical effective temperatures, $T_{\rm eff}$, and bolometric corrections, $B.C.$, from a combination of satellite absolute fluxes and the Narrabri intensity-interferometer angular diameters. In most cases, however, one must rely on photometric indices such as c_0, calibrated by means of the above-mentioned empirical $T_{\rm eff}$ and $B.C.$ data.

Using these empirical data, Davis and Shobbrook (1977) derived mean relationships between c_0, $T_{\rm eff}$ and $B.C.$ for the $T_{\rm eff}$ range from 11000 to 34000 K, separately for luminosity classes III-V and Ia. According to Shobbrook (1979b), for the region of the β Cephei stars the mean relationships give:

$$\log T_{\rm eff} = -0.640\, c_0 + 4.405 \qquad (3)$$

$$B.C. = 3.26\, c_0 - 2.52 \qquad (4)$$

Using the same data, and $\theta_{\rm eff} = 5040/T_{\rm eff}$ instead of log $T_{\rm eff}$, Sterken and Jerzykiewicz (1980) found:

$$\theta_{\rm eff} = 0.282\, c_0 + 0.203 \qquad (5)$$

The above equations are based on the assumption that for B stars of luminosity classes III-V the correlation between bolometric flux and the size of Balmer jump is independent of surface gravity. This assumption was not made by Balona (1984), who calibrated $\log T_{\rm eff}$ and $B.C.$ in terms of c_0 and β, using β as a measure of surface gravity. Balona (1984) noted that the number of stars for which empirical values had been derived by Code et al. (1976) is too small to allow an adequate two-dimensional calibration. He solved the problem by using Kurucz's (1979) model atmospheres to establish the functional form of the calibrations, and the empirical values, to fix the zero points. All his calculations were restricted to models with solar abundances.

Balona's calibration equations are the following:

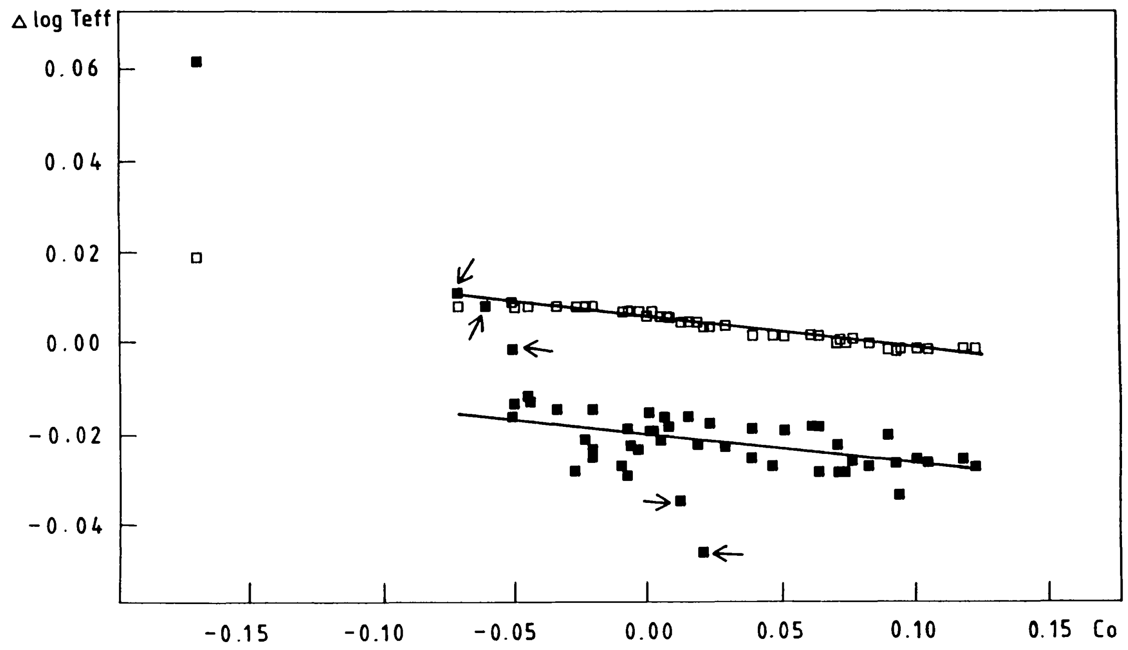

Fig. 15. A comparison of the effective temperatures of the β Cephei stars, obtained from the calibrations of (1) Davis and Shobbrook (1977), (2) Sterken and Jerzykiewicz (1980), and (3) Balona (1984). Open squares represent the differences "$\log T_{\rm eff}(1)$ *minus* $\log T_{\rm eff}(2)$", while filled squares, the differences "$\log T_{\rm eff}(3)$ *minus* $\log T_{\rm eff}(2)$". The leftmost squares correspond to V986 Oph. The two straight lines were fitted to the $\log T_{\rm eff}$ differences in the interval $-0.07 < c_0 < 0.13$ by the method of least squares. In the case of the lower straight line, the five deviant points (indicated by arrows) were not used in the fit. These points correspond to, from left to right, NGC 6231-110, NGC 6231-238, NGC 6231-150, 19 Mon, and NGC 3293-65.

$$\log T_{\rm eff} = 3.9036 - 0.4816\,[c] - 0.5290\,[\beta] - 0.1260\,[c]^2 + 0.0924\,[\beta]\,[c] - 0.4013\,[\beta]^2 \quad (6)$$

$$B.C. = 0.2900 + 2.8467\,[c] + 2.8334\,[\beta] + 0.6481\,[c]^2 - 0.2997\,[\beta]\,[c] + 2.1487\,[\beta]^2 \quad (7)$$

where $[\beta] = log(\beta - 2.500)$, and $[c] = log(c_0 + 0.200)$.

For β Cephei stars of Table I, a comparison of the effective temperatures, obtained using the above-mentioned three calibrations, is presented in Figure 15[3]. As could be expected, the agreement between the one-dimensional calibrations is quite good: the difference between the effective temperatures derived from the mean relation of Davis and Shobbrook (1977) and those computed from eq. (5) varies from zero around $c_0 = 0.07$ to an insignificant 500 K at $c_0 = -0.07$. On the other hand, except for a few deviant values, Balona's (1984) effective

[3] In this and the following three figures, 27 (EM) CMa is not plotted because its β index is strongly affected by emission (see Section 3.3).

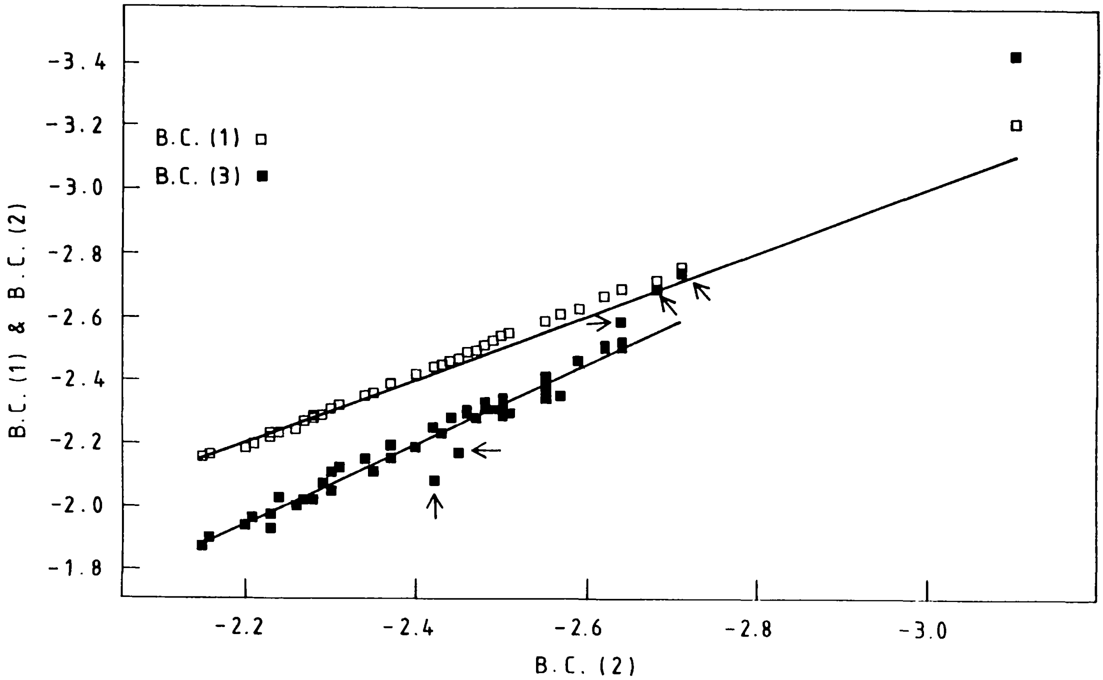

Fig. 16. A comparison of the bolometric corrections of the β Cephei stars, obtained from the calibrations of (1) Davis and Shobbrook (1977), (2) Sterken and Jerzykiewicz (1980), and (3) Balona (1984). The rightmost squares correspond to V986 Oph. The long straight line-segment has unit slope and zero intercept. The short one was fitted to $B.C.$(3) by the method of least squares, omitting V986 Oph and the five stars mentioned in the caption to Figure 15 (they are indicated by arrows).

temperatures are systematically lower by 0.026 in the logarithm from those of Davis and Shobbrook (1977). This corresponds to about 1600 K and 1200 K at the hot and cool ends of the $-0.07 < c_0 < 0.13$ interval, respectively.

Bolometric corrections of β Cephei stars are used as coordinates in Figure 16. The slight deviation of Davis and Shobbrook's (1977) bolometric corrections from the straight line of unit slope and zero intercept, seen in this figure, is caused by the small systematic difference between their and Sterken and Jerzykiewicz's $T_{\rm eff}$ scales. However, the systematic difference between Balona's (1984) $B.C.$ scale and that of Davis and Shobbrook (1977) is somewhat greater than would be expected from the above-mentioned systematic difference of 0.026 in the logarithm between their $T_{\rm eff}$ scales. This fact is illustrated in Figure 17, where the actual systematic difference between the $B.C.$ scales is represented by the solid line, while that resulting from the difference in the $T_{\rm eff}$ scales, is given by a dashed line.

4.2. The Absolute Magnitudes

There are two independent lines of evidence showing that for the early B stars the well-known Crawford (1978) calibration of the β index may give absolute magnitudes too bright by about half a magnitude. Firstly, in the case of the primary components of Spica and 16 Lac, for which the pulsation constants can be derived directly because their radii and masses were estimated from the orbital solutions, the pulsation constants obtained from c_0 and β are significantly smaller than the directly derived ones (see the next section). Jerzykiewicz and Sterken (1980) pointed out that this discrepancy would be removed if the absolute magnitudes of Spica and 16 Lac, obtained from β via Crawford's (1978) calibration, were increased by $0\overset{m}{.}5$. Secondly, Shobbrook (1983b) has found that Crawford's (1978) calibration applied to the brightest NGC 3293 stars – including the ten β Cephei variables – indicates a distance modulus about $0\overset{m}{.}6$ larger than that derived by means of main-sequence fitting to the fainter stars of the cluster. According to Shobbrook (1983a), a linear relation which fits the brightest members of NGC 3293 is:

$$M_V = 51.08\,\beta - 136.4 \tag{8}$$

This preliminary result, applicable only over the spectral-type range B0-B2, was soon superseded by a calibration derived by Balona and Shobbrook (1984) from $uvby\beta$ data on 421 stars in 13 open clusters. This new calibration, valid over the entire B range, has the following form:

$$M_V = a_0 + a_1 \log(\beta - 2.515) + a_2\,[g] + a_3\,[g]^3 \tag{9}$$

with $[g] = log(\beta - 2.515) - 1.60\,(c_0 + 0.322)$, $a_0 = 3.4994$, $a_1 = 7.2026$, $a_2 = -2.3192$, and $a_3 = 2.9375$. It fits the data with a standard error of $0\overset{m}{.}43$ per star.

For the β Cephei stars, a comparison of the three M_V scales is presented in Figure 18. In the interval $-4\overset{m}{.}0 > M_V > -5\overset{m}{.}0$, the new absolute magnitudes agree with each other and are about $0\overset{m}{.}5$ fainter than Crawford's (1978). For $M_V > -4\overset{m}{.}0$, Shobbrook's (1983a) M_V are about $0\overset{m}{.}6$ fainter than Crawford's (1978), but the difference between the scale of Balona and Shobbrook (1984) and that of Crawford (1978) decreases to about $0\overset{m}{.}2$ at $M_V = -2\overset{m}{.}8$. For the primary components of Spica and 16 Lac, the absolute magnitudes computed from eq. (9) are, respectively, $0\overset{m}{.}4$ and $0\overset{m}{.}3$ fainter than those obtained by means of the calibration of Crawford (1978).

The effective temperatures and bolometric absolute magnitudes of the β Cephei stars, derived from the c_0 and β indices of Table IIA by means of eqs. (6), (7) and (9), are given in Table IVA of the Appendix.

The absolute magnitudes of the β Cephei variables that are members of open clusters can also be obtained from their V magnitudes, corrected for the interstellar absorption, and the mean distance moduli of the clusters. Such M_V of the NGC 3293, NGC 4755, and NGC 6231 β Cephei stars have been derived by Shobbrook (1985). In most cases, these values agree to better than the above-mentioned standard error of $0\overset{m}{.}43$ with M_V obtained from β and c_0 via the calibration of Balona and Shobbrook (1984). The exceptions are represented in Figure 18 by means of filled squares with arrows; ordinates of the arrowheads indicate M_V derived from V and the cluster moduli. The downward pointing arrows suggest

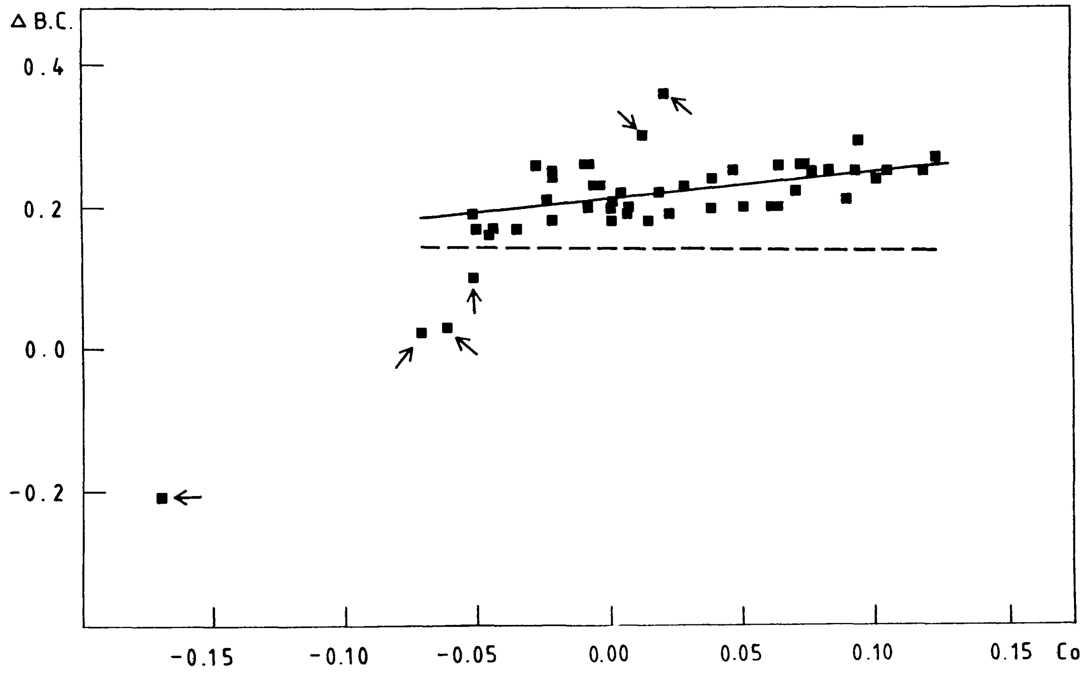

Fig. 17. Differences between the bolometric corrections of the β Cephei stars, derived from the calibrations of Balona (1984) and Davis and Shobbrook (1977), plotted against c_0. The solid straight line was fitted to the differences by the method of least squares, omitting the same six stars as in Figure 16 (indicated by arrows). The dashed line represents the $B.C.$ difference, equal to 0.14 mag, resulting from the systematic difference of 0.026 in the log $T_{\rm eff}$ scales of Balona (1984) and Davis and Shobbrook (1977).

emission at Hβ. For NGC 6231-110, 150, and 253, this possibility has already been noted by Shobbrook (1985). On the other hand, the $0^{\rm m}.5$ discrepancy in the case of NGC 3293-27 (the only upward pointing arrow) indicates, according to Shobbrook (1985), that the star is an equal-component binary.

4.3. The Evolutionary State of Field β Cephei Stars from Their Position in the Theoretical H-R Diagram

Kopylov (1959) and Schmalberger (1960) were the first to compare the position of β Cephei stars in the theoretical H-R diagram with evolutionary tracks. Both these authors noted that the β Cephei sequence runs parallel to and significantly above the theoretical ZAMS. Moreover, Kopylov (1959) concluded that all early B stars go through a relatively short-lived β Cephei stage during their main-sequence evolution, whereas Schmalberger (1960) suggested that there may be a close correlation between β Cephei-type variability and either the terminal phases of main-sequence evolution or "the structural reorganization which seems likely to follow." In modern terms, this places the β Cephei variables in one of the following evolutionary stages: close to the end of core hydrogen-burning, the secondary contraction, or the shell hydrogen-burning.

Subsequent work by Percy (1970b), Watson (1972), Lesh and Aizenman (1973a),

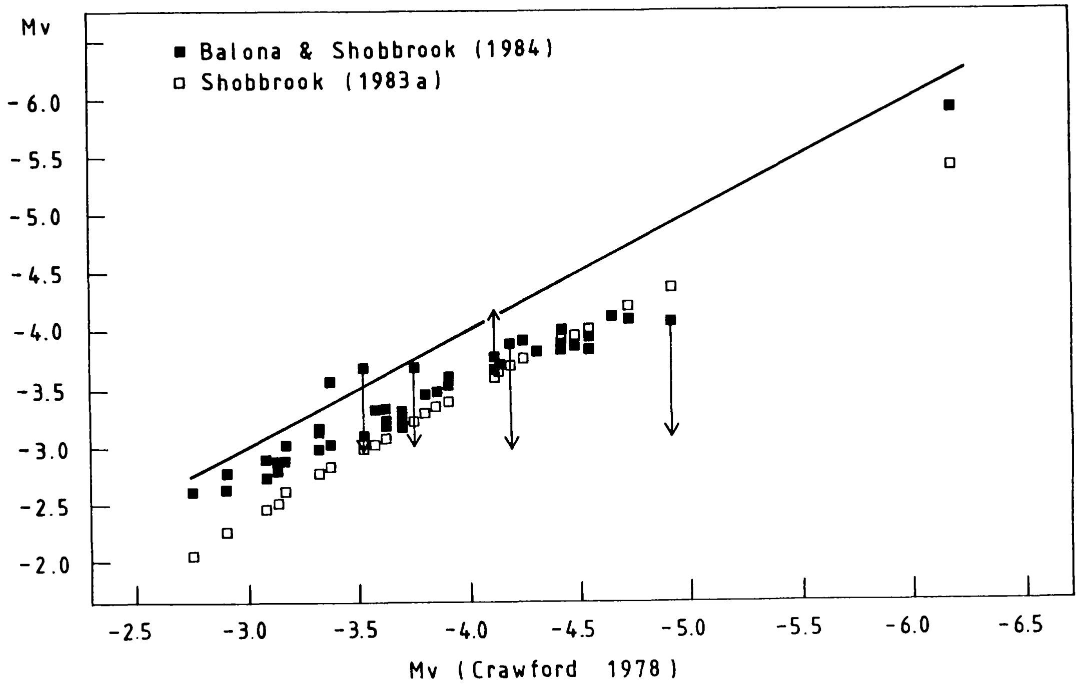

Fig. 18. The absolute visual magnitudes of the β Cephei stars, obtained from the calibrations of Shobbrook (1983a) and Balona and Shobbrook (1984), plotted against the absolute visual magnitudes from the calibration of Crawford (1978). The diagonal line has unit slope and zero intercept. The rightmost squares correspond to V986 Oph. Filled squares with arrows represent, from left to right, the open-cluster variables, NGC 6231-110, NGC 6231-150, NGC 3293-27, NGC 6231-253, and NGC 3293-65. The ordinates of the arrowheads indicate absolute magnitudes, derived from the cluster distance-moduli.

and Shobbrook (1978b) led to a general agreement that Schmalberger's (1960) conclusion was essentially correct as far as the position of β Cephei stars in the theoretical H-R diagram is concerned. Typical in this respect is the result of Lesh and Aizenman (1973a), reproduced here in Figure 19, where one can see the coincidence of the β Cephei instability strip with the S-band region of theoretical evolutionary tracks, where the above-mentioned evolutionary stages are occurring.

There was no agreement, however, about the evolutionary state of these stars. Watson (1972) believed that the majority of β Cephei variables are in the late core hydrogen-burning phases. On the other hand, Lesh and Aizenman (1973a) suggested that the variables are in one of the two later stages of evolution, while nonvariable B stars are core hydrogen-burning objects. Both these conclusions were based on considerations of theoretical evolutionary lifetimes and observed numbers of β Cephei stars. The disagreement was caused by the difference in estimating the relative frequency of the variable and constant stars in the instability strip, mentioned in Section 3.6.

Shobbrook's (1978b) investigation of the position of β Cephei stars in the theoretical H-R diagram and of their relative frequency in the instability strip, based on his extensive β photometry and published *uvby* data (see Section 3.3), led him to believe that these variables must be very near the end of core hydrogen-

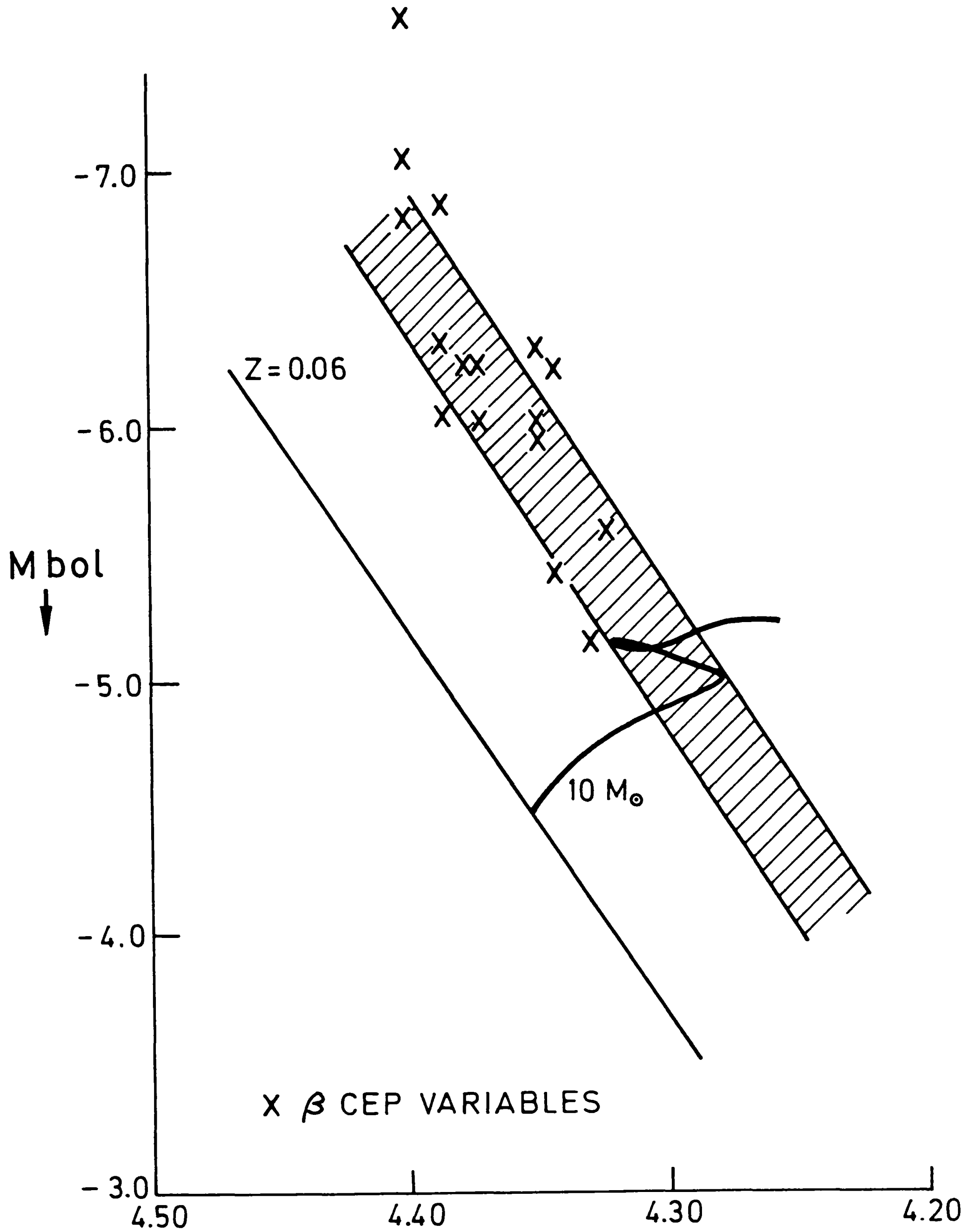

Fig. 19. The location of the β Cephei stars in the theoretical H-R diagram according to Lesh and Aizenman (1973a).

burning. On the other hand, Jerzykiewicz and Sterken (1980) maintained that the evolutionary state of these stars may span the range from about half-way through the main-sequence phase to beyond the end of core hydrogen-burning.

The theoretical H-R diagram on which the latter conclusion was based is reproduced in Figure 20. The β Cephei variables (filled circles), constant stars (open circles), the ridge line of the instability strip (rightmost solid line), its boundaries (dashed lines), and Crawford's (1978) ZAMS were transformed from the $c_0 - \beta$ diagram, shown in Figure 7. V986 Oph ($\log T_{\rm eff} = 4.506$ and $M_{bol} = -9\overset{\rm m}{.}25$) is

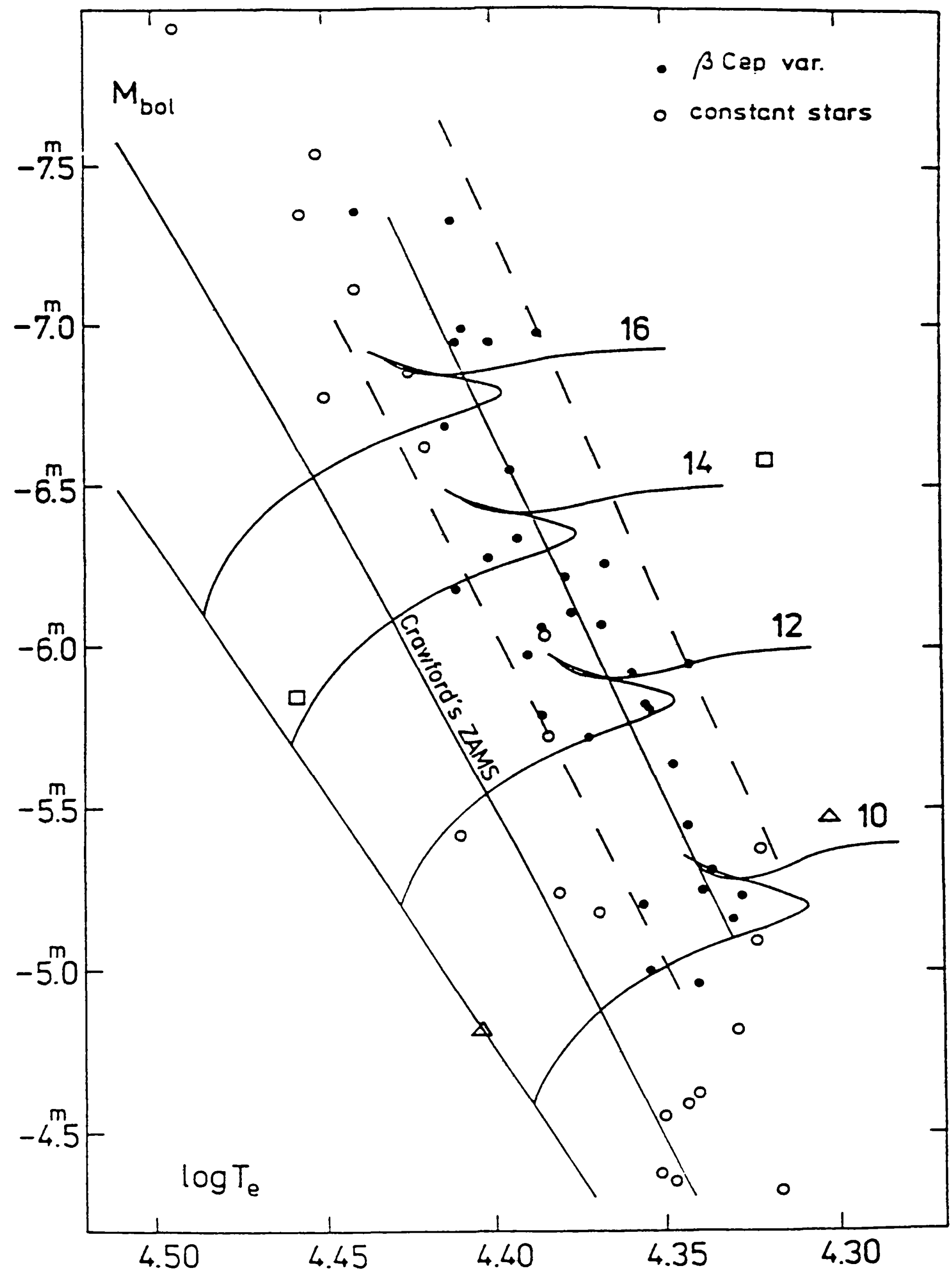

Fig. 20. The β Cephei variables (filled circles) and constant early B stars (open circles) in the theoretical H-R diagram (from Sterken and Jerzykiewicz 1980). See Section 4.1 for explanation of the remaining symbols.

not plotted. The c_0-to-effective-temperature transformation and bolometric corrections were based on the OAO-2 empirical scales of Code et al. (1976). In particular, $\log T_{\rm eff}$ was derived from eq. (5). The absolute visual magnitudes were obtained from β by means of the calibration of Crawford (1978).

The evolutionary tracks of 10, 12, 14, and 16 $M_\odot$ models, shown in Figure 20, have been computed by de Loore et al. (1978), using a modified Paczynski stellar evolution program, Cox-Steward opacity tables, and initial chemical composition $X = 0.70$ and $Z = 0.03$. The leftmost open triangle and open square represent the zero-age models of Stothers (1976) for M equal to 10.9 and 15 M⊙ respectively. In these models, Carson's radiative opacities were used and the composition was $X = 0.73$, $Z = 0.02$. The rightmost open triangle and open square indicate the position of these models near the end of the core hydrogen-burning phase.

As can be seen from Figure 20, the zero-age models fall considerably below the observed zero-age main-sequence. This well-known discrepancy makes any conclusion concerning the evolutionary state of the field β Cephei stars somewhat uncertain. Lesh and Aizenman (1973a) removed the discrepancy by suitably increasing the model opacities. Using Z as an "opacity parameter", they achieved agreement between the zero-age models and observed ZAMS for $Z = 0.06$. The theoretical ZAMS and evolutionary track in Figure 19 were computed with this value of Z. However, Sterken and Jerzykiewicz (1980) pointed out that if the same procedure were applied to Figure 20, most β Cephei stars would probably be found in the core hydrogen-burning phase. In addition, the terminal-age main sequence would then approximately coincide with the upper boundary of the β Cephei instability strip.

Investigation of β Cephei stars in NGC 3293 and NGC 6231 by Balona and his co-workers, reviewed in Section 3.5, led Balona (1987a) to the conclusion that the β Cephei phenomenon is not confined to the S-band region of the evolutionary tracks, as thought by Schmalberger (1960) and his followers, but occurs from the ZAMS until the end of core hydrogen-burning. This conclusion was reached without a direct comparison with the theoretical evolutionary tracks, and therefore the above-mentioned discrepancy between the theoretical and observed ZAMS did not emerge. Note, however, that the possibility of the phenomenon persisting beyond the end of the core hydrogen-burning phase has not been excluded, because stars in the later two evolutionary stages of the S-band region are simply not found in NGC 3293 and NGC 6231.

4.4. Secular Period Changes

Detecting evolutionary changes in β Cephei stars was first considered by Struve (1955a). He expressed the idea in the following words:

> "... if a B-type star evolves with constant mass in such a manner as to describe an evolutionary track away from the main sequence, its mean density would decrease. If such a star should pulsate in accordance with the relation $P\sqrt{\rho}$ = constant, the corresponding change in the period may prove to be detectable."

The long-term behaviour of the period of a variable star can be studied by means of an $O - C$ diagram, showing the difference between the observed epoch of maximum (or minimum) light, O, and the epoch predicted with an assumed period, C, as a function of the cycle number, E. An example, based on all available epochs of maximum light of δ Cet and an assumed period of $0\overset{d}{.}16113735$, is displayed in Figure 21. In this figure, the straight line corresponds to a constant

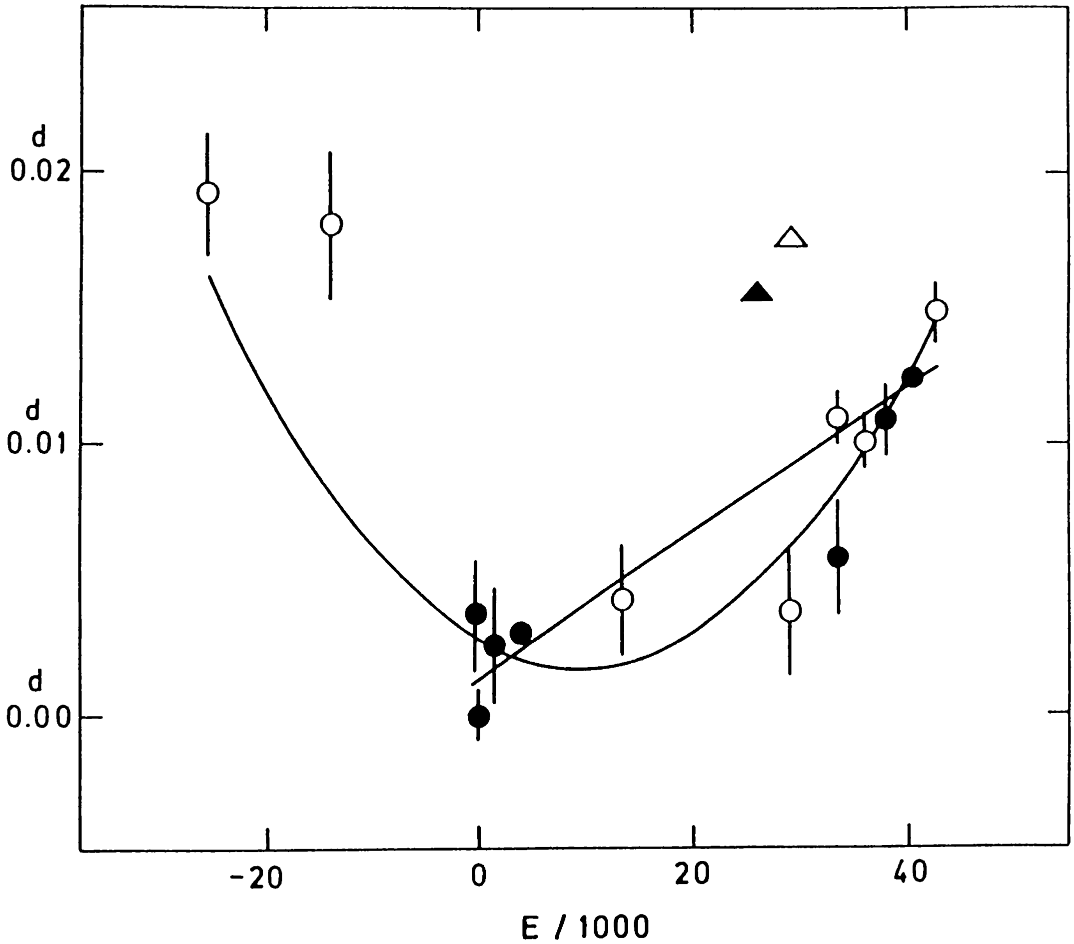

Fig. 21. An O-C diagram for δ Cet. Open symbols represent O-C's derived from observations on one or two nights, whereas filled symbols, those from observations on four or more nights. Reproduced from Jerzykiewicz *et al.* (1988).

period of $0\overset{d}{.}16113762$, while the parabola corresponds to a period increasing with a constant rate equal to 0.47 ± 0.09 sec/century.

Eggleton and Percy (1973) calculated the rate of period change along an evolutionary track of a massive B star and compared it with observed rates of period change in eight β Cephei variables. They concluded that the uncertainty in the masses of these stars and the possibility of nonevolutionary period changes make it impossible to determine precisely their evolutionary state. Nevertheless, they regarded the secondary gravitational contraction or early shell hydrogen-burning phases of evolution as unlikely. On the other hand, a similar investigation led Lesh and Aizenman (1974) to the conclusion that it would be difficult to reconcile the majority of the observed period changes with the core hydrogen-burning phase, unless large nonevolutionary effects were present.

The hypothesis that β Cephei stars showing period changes are hydrogen-burning objects in which nonevolutionary processes affect the periods has been advanced by Chapellier (1985). He maintains that all O-C diagrams can be better represented with polygonals than with parabolae. In other words, the periods always change abruptly after intervals of constancy, rather than that they decrease or increase with a constant rate. The stars for which Chapellier (1985) finds largest abrupt period changes include BW Vul, β Cep, and σ Sco. According to Chapellier (1985), the mechanism of these abrupt period changes may be similar to that suggested by Sweigart and Renzini (1979) for the RR Lyrae stars. This mechanism links period changes with the composition changes, caused by discrete mixing events in a semiconvective zone. As an argument in favour of this hypothesis, Chapellier (1985) regards the fact that β Cephei stars follow a correlation between abrupt period change and period. This correlation, discovered by Hoffleit (1976), is apparently obeyed by various kinds of pulsating variables, ranging from RR Lyrae stars to Miras. Chapellier (1985) notes that β Cephei stars fall close to the short-period projection of the correlation and concludes that the observed period changes are "accidental events without any evolutionary significance".

Chapellier's (1985) work can be criticized on two accounts. Firstly, it is not exactly true that period changes in β Cephei stars are always abrupt. In the case of β Cep, for example, the observed period changes are smooth and have been recently shown by Pigulski and Boratyn (1992) to be caused by the light-time effect in a long-period, eccentric binary orbit[4]. Secondly, it was demonstrated by Jerzykiewicz (1986) that Hoffleit's (1976) correlation has arisen as a result of observational selection. Consequently, no conclusions about the true nature of period changes should be based on it.

In β Cep, the period changes unaccounted for by the light-time effect are below 0.1 sec/century (Pigulski and Boratyn 1992), indicating that the star is indeed a hydrogen-burning object. The light-time effect has also been found in BW Vul (Odell 1984, Jiang 1985, and Pigulski 1992b) and σ Sco (Pigulski 1992a). In both these cases, however, the effect is accompanied by secular period increases of the order of 3 sec/century.

The mass of BW Vul, estimated from its position in the theoretical H-R diagram (Figure 19 or 20), is equal to 13 $M_\odot$, and its pulsation constant (see the next section) is equal to 0.037. Using these numbers and the theoretical rates of period change given by Eggleton and Percy (1973) and by Lesh and Aizenman (1974), Sterken and Jerzykiewicz (1990b) concluded that if the above-mentioned period increase of 3 sec/century has an evolutionary origin, it can be consistent only with the early phases of shell hydrogen-burning[5]. Thus, BW Vul may be more evolved than the majority of β Cephei stars.

As a second argument in favour of this idea, Sterken and Jerzykiewicz (1990b) use the fact that the star has exceptionally large amplitudes of light and radial-velocity variations. The argument involves the theory of limiting the growth of pulsation amplitudes. Dziembowski (1988) has recently suggested that the growth of pulsation amplitudes in β Cephei stars is limited by resonant mode coupling,

[4] That the light-time effect may be responsible for part of the observed period changes in β Cephei and in σ Sco was for the first time suggested by Chapellier(1990)

[5] A similar possibility was considered by Percy (1971).

the same mechanism that is presumably operating in δ Scuti stars. In the ZAMS δ Scuti models this mechanism is so effective that it may halt the growth of pulsation amplitudes before they reach detectability threshold. In this way, Dziembowski and Krolikowska (1985) have accounted for the large percentage of constant stars in the lower cepheid instability strip. The large amplitudes of the more luminous δ Scuti stars indicate that the mechanism is less effective for evolved objects. BW Vul may be a β Cephei analogue of these evolved, large-amplitude δ Scuti stars.

As pointed out by Pigulski (1992a), similar arguments apply to σ Sco, which may therefore be the second β Cephei star in the early phases of shell hydrogen-burning.

As shown by Pigulski (1992b), all other β Cephei stars for which existing data permit a reliable analysis of secular period changes are γ Peg, δ Cet, and ξ^1 CMa. According to Pigulski (1992b), the analysis yields secular period increases amounting to 0.06 ± 0.02, 0.28 ± 0.02, and 0.37 ± 0.05 sec/century, respectively, all consistent with the phase of core hydrogen-burning. Thus, by evidence of the secular period changes, these three stars and β Cep, mentioned earlier in this section, are main-sequence objects. Note that they all have much smaller light and radial-velocity amplitudes than BW Vul and σ Sco.

An important outcome of Pigulski's (1992b) analysis is that, apart from the apparent variations of the period due to the light-time effect, he found not a single case of non-evolutionary period changes. An obvious possibility is to look for such changes in objects of known evolutionary state, for example the β Cephei stars in NGC 3293. The first attempt to investigate the secular stability of the pulsation period of HD 92024 = NGC 3293-5, undertaken by Jerzykiewicz and Sterken (1992), was inconclusive because the time-base covered by observations was too short.

5. Periods and pulsation modes

5.1. Multiperiodicity and Nonradial Pulsations

Some β Cephei stars are strictly periodic, but most show long-period variations of the amplitudes of their radial-velocity and light curves. As we mentioned in the Introduction, in the case of β CMa this phenomenon was studied by Meyer (1934), who found that the star's radial-velocity amplitude varied with a period of $49^{\mathrm{d}}.1$. Meyer (1934) was able to explain his observations in terms of "beating" between two sine-curves with slightly different short periods, $P_1 = 6^h0^m$ and $P_2 = 6^h2^m$. The reality of this decomposition of the observed radial-velocity curve into two short-period components has been supported by the fact that P_2 was identical with the period of the variation of the widths of spectral lines, discovered by Henroteau (1918).

An attempt to account for the "beat phenomenon" and the line-profile variations in β CMa was made by Ledoux (1951). He showed that closely-spaced frequencies arise in a nonradially pulsating star as a result of slow rotation. In the linear

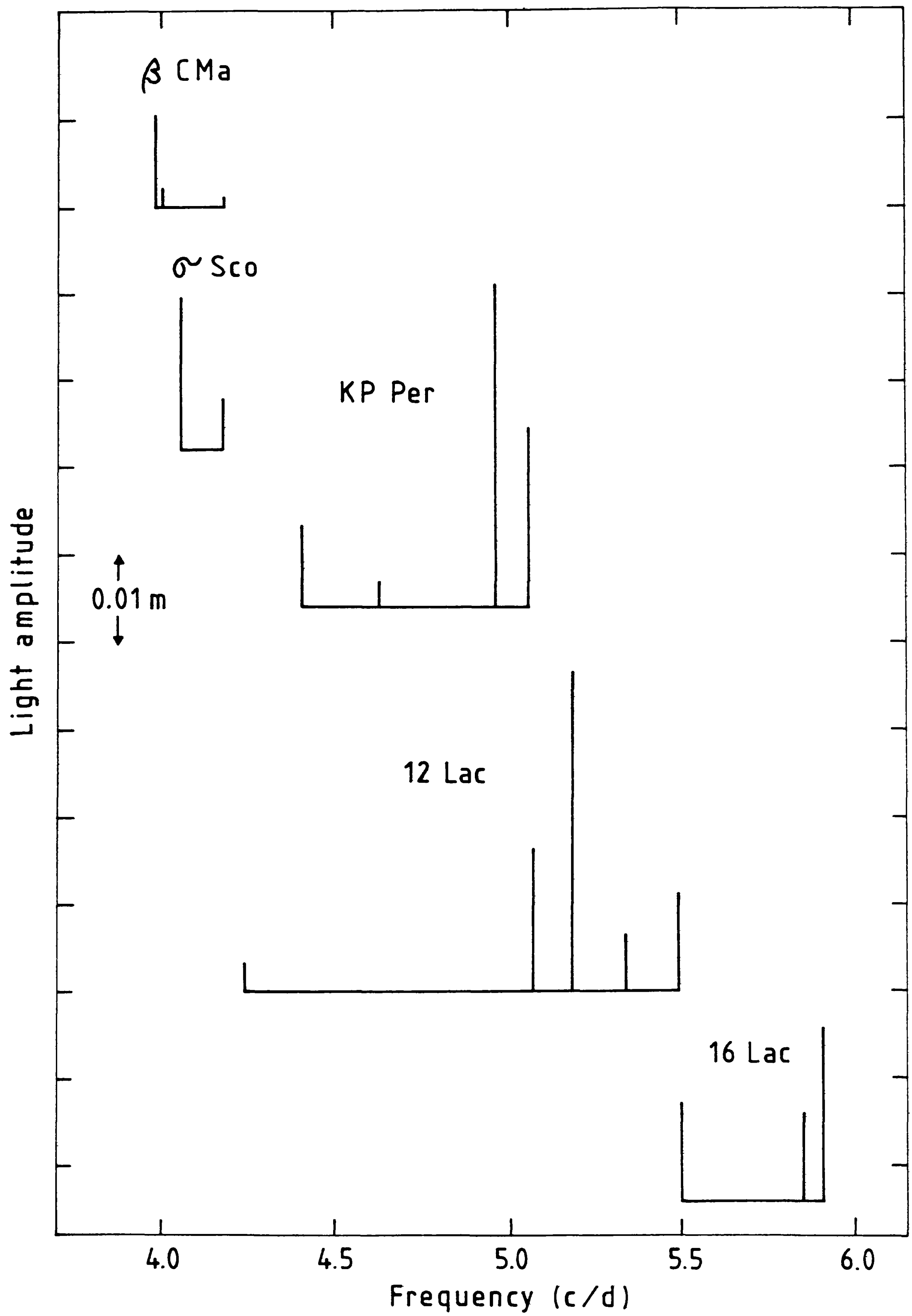

Fig. 22. Light amplitudes as a function of frequency (cycle per day) for β CMa (Shobbrook 1973a), σ Sco (Jerzykiewicz and Sterken 1984), KP Per (Jerzykiewicz and Mysior 1980), 12 Lac (Jerzykiewicz 1978), and 16 Lac (Jarzębowski *et al.* 1979). Harmonics and cross-terms are not shown (Figure 1 of Jerzykiewicz and Sterken 1980).

approximation, the surface of a star pulsating in a nonradial mode has the shape of a spherical harmonic of degree l and order m, multiplied by a time-dependent factor of the form $\exp(2\pi i \omega_{l,m} t)$, where $\omega_{l,m}$ is the pulsation frequency. For an observer at rest, the pulsation frequency satisfies the following equation:

$$2\pi \omega_{l,m} = 2\pi \omega_l - m(1-C)\Omega \qquad -l \leq m \leq l, \tag{10}$$

which is accurate to first order in Ω, the angular velocity of the star's rotation. The quantity C is a function of l and the equilibrium structure of the star. If all m modes belonging to a given l were excited, there would arise $2l+1$ equidistant frequencies, distributed symmetrically around ω_l. The latter frequency corresponds to a stationary pulsation, symmetric with respect to the axis of rotation, while $\omega_{l,m}$ correspond to waves traveling in the same direction as rotation if $m < 0$, or in the opposite direction if $m > 0$.

Ledoux (1951) also demonstrated that the $l = 2$, $m = \pm 2$ modes produce line profile and radial-velocity variations with a period $P = 1/\omega_{l,m}$ and a phase difference equal to $\pm\pi/2$. The effect is greatest when the star is seen equator-on, and disappears for the aspect angle, i, equal to zero, that is, when the axis of rotation is parallel to the line of sight. The stationary pulsation, $l = 2$, $m = 0$, turned out to produce radial-velocity variations only, unless i was close to 54°. For this aspect angle, the velocity was constant and the line profiles were variable, but with a period equal to $P/2$.

Making use of these results, Ledoux (1951) identified the observed P_2-variation in β CMa with an $l = 2, m = 2$ traveling wave, and the P_1 variation with the stationary pulsation, $l = 2, m = 0$, assuming, moreover, that the line of sight is close to the equatorial plane. For the f mode of the standard model this gave a period of the star's rotation of the order of 80 days, and the equatorial velocity of rotation, V_e, equal to about 8 to 10 km s^{-1}. However, the sign of the above-mentioned $\pi/2$ phase-shift was wrong, that is, the computed lines were broadest on the ascending branch of the P_2 velocity variation, and not on the descending branch, as observed. Of course, a wave traveling in the same direction as rotation would yield the correct phase relation, but then the broadening would be associated with the shorter period, contrary to what is observed. In addition, it was later found by McNamara and Hansen (1961) that for β CMa the observed $V_e \sin i = 28\, km\, s^{-1}$, that is, about three times the predicted V_e.

In spite of these discrepancies, the possibility of explaining the beat phenomenon and the line-profile variations in terms of the $l = 2$ pulsations in the presence of slow rotation was further investigated by Osaki (1971). He confirmed the result of Ledoux (1951) that a superposition of a stationary pulsation and a traveling wave results in a long-period modulation of the velocity amplitude and a line-profile variation, the latter caused mainly by the traveling wave. Osaki (1971) also showed that rotational velocities, obtained for several β Cephei variables from the observed beat periods via eq. (10) under the assumption that the $m = -2$ and

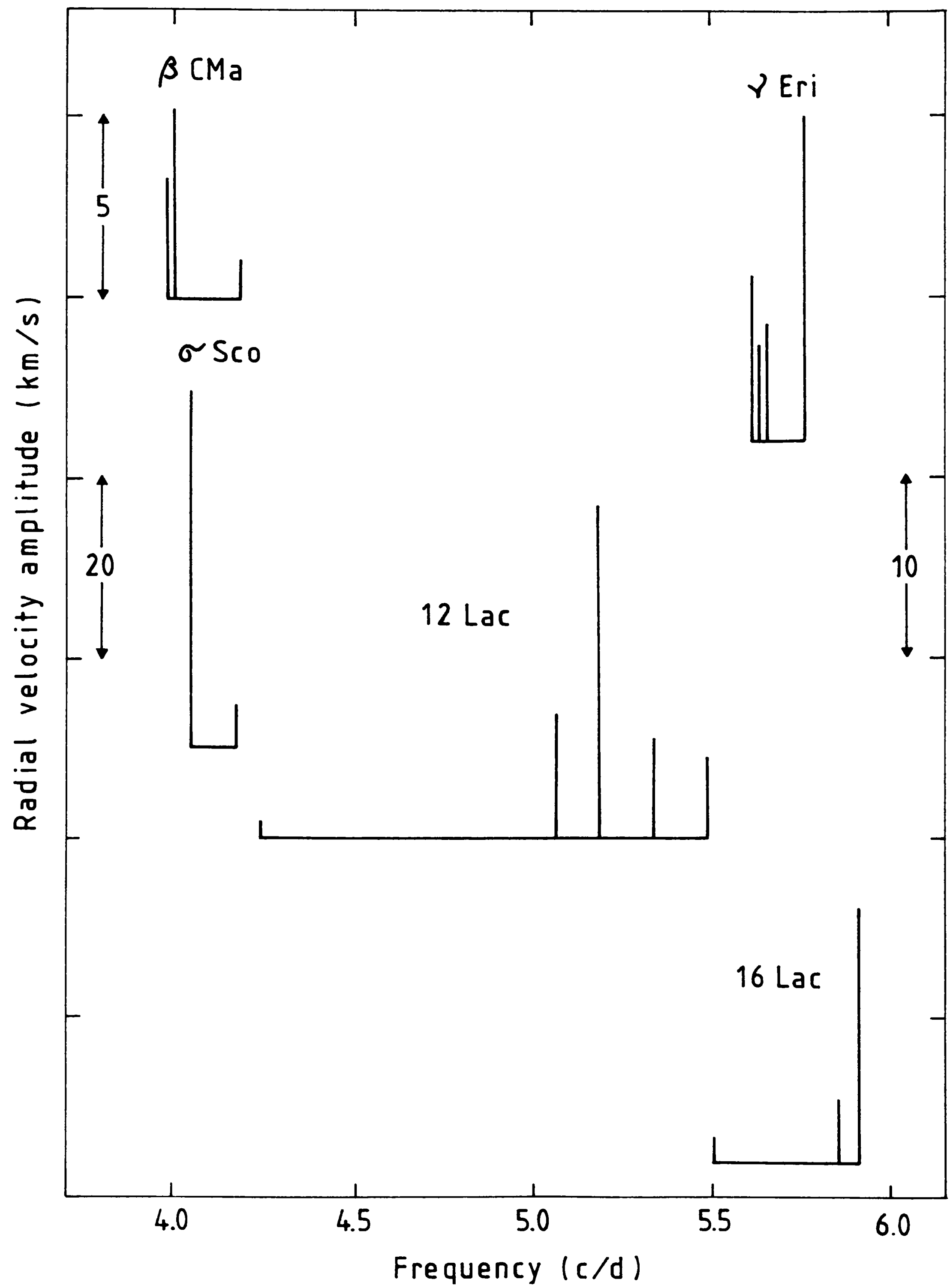

Fig. 23. The radial-velocity amplitudes as a function of frequency for β CMa (Shobbrook 1973a), ν Eri (Kubiak 1980), σ Sco (Kubiak 1980), 12 Lac (Jerzykiewicz 1978), and 16 Lac (Fitch 1969). Harmonics and cross-terms are not shown.

$m = 0$ pulsations were both excited, are consistent with the measured $V_e \sin i$ values for i close to 90°.

Due to the work of Fitch (1969), Shobbrook (1973a) and others, it is now clear

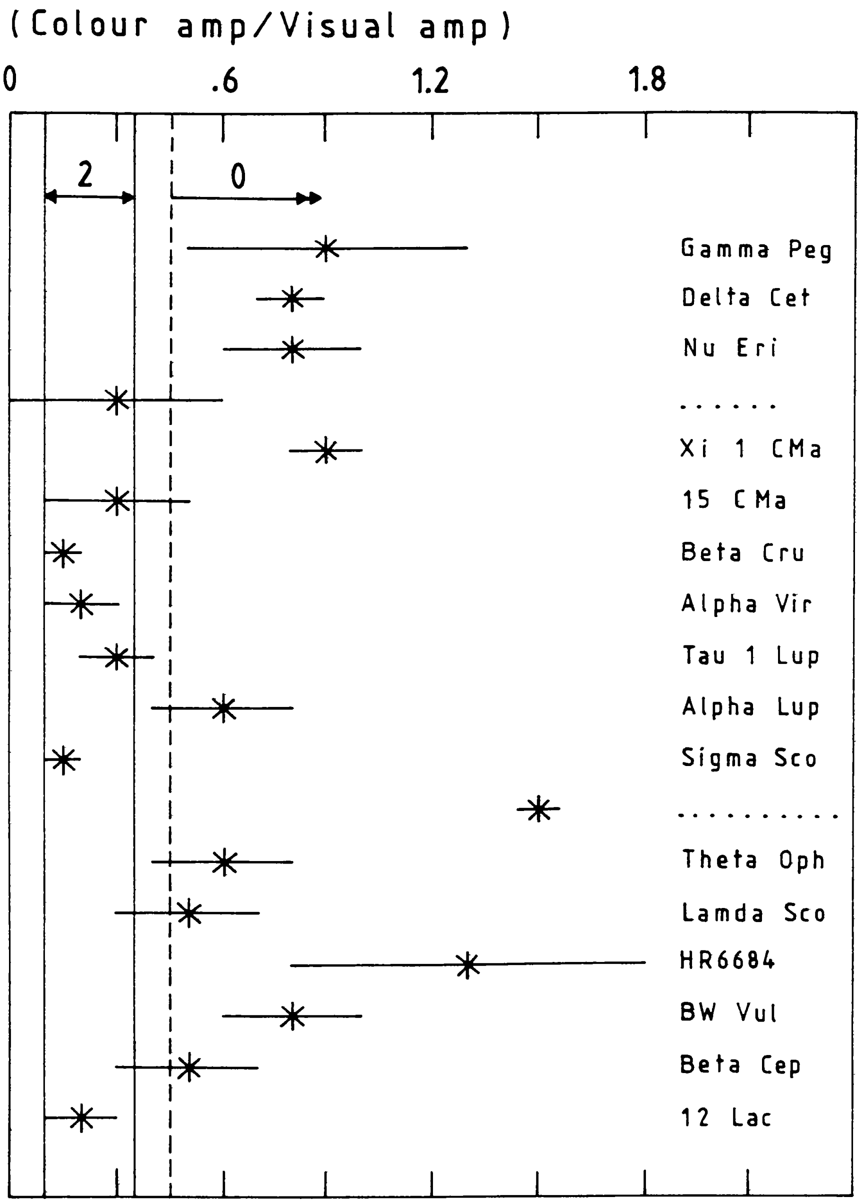

Fig. 24. A_{U-V}/A_V data for β Cephei stars, with radial pulsation ($\geq$ 0.45) and $l = 2$ pulsation (0.1 to 0.35) guidelines shown. The $l = 1$ pulsations occupy a small intermediate region. From Watson (1988).

that most "beat" β Cephei variables are multiperiodic. This fact is illustrated in Figs. 22 and 23, reproduced from Jerzykiewicz and Sterken (1980), where the light and radial-velocity frequency spectra of several large-amplitude variables are shown.

For each of the four stars common to Figs. 22 and 23 (that is, β CMa, σ Sco, 12 Lac, and 16 Lac) the frequencies in the light and radial velocity match one another. This applies to all well-observed β Cephei variables, including the singly-periodic ones. As we have seen in Section 4.4, secular frequency-changes are very small. However, large-amplitude changes on a time scale of years occur in some stars, for example, in 16 Lac (Fitch 1969, Jerzykiewicz 1976, Jarzębowski et al. 1979). The most striking and well-known case is that of Spica, discussed already in Section 3.6.

In 12 Lac and ν Eri equally-spaced frequency triplets are seen. They can probably be explained in terms of first order rotational splitting of a nonradial mode. However, rotational splitting alone does not account for other frequencies occurring in these two stars, and equally-spaced frequency patterns were not found in β CMa, KP Per, and 16 Lac. Ledoux's (1951) and Osaki's (1971) hypothesis of the $l = 2$ pulsations is inadequate in these cases. One is forced to conclude that in a multiperiodic β Cephei star different l modes can be simultaneously present.

5.2. Identification of l and m

Several attempts have been made to identify the spherical harmonic degree, l, and order, m, of the pulsation modes occurring in β Cephei stars. Three methods proved to be particularly useful. The first, a generalization of the approach of Ledoux (1951) and Osaki (1971), consists in computing sequences of spectral line profiles for a number of l, m, V_e, and i combinations, and matching the computed line profiles with the observed ones. The second, developed by Stamford and Watson (1977), exploits the wavelength dependence of the pulsation amplitude as a discriminant of l. The latter method is insensitive to m. A third method was elaborated by Balona (1986a) and is based on the first two or three moments of observed line profiles. Mode identification is then done by obtaining the phases and amplitudes of the various oscillations from a periodogram analysis of the moments. In a subsequent paper, Balona (1986b) describes a complementary method that is applicable when the pulsational velocity amplitude is considerably smaller than the projected rotational velocity.

The first method was used by Kubiak (1978, 1980b), Campos and Smith (1980), and Smith (1980). Kubiak based his work on published line-profile observations, mainly photographic ones, whereas Campos and Smith utilized Reticon observations of the Si III λ4567 line. According to Campos and Smith (1980), only a radial mode can produce the variable line-profiles observed in the singly-periodic stars γ Peg, β Cep, and δ Cet, as well as in the multiperiodic star σ Sco. However, Kubiak (1978, 1980b) was able to account for the line-profile variations of β Cep and σ Sco using nonradial modes. There is no satisfactory explanation for these discrepancies. Kubiak's results are summarized in Table III, reproduced from Jerzykiewicz and Sterken (1980). This table also contains Smith's (1980) l and m identifications for 12 Lac.

The second method has recently been applied by Watson (1988) to a number of β Cephei stars for which the ultraviolet and visual amplitudes are known. The results, summarized in Figure 24, indicate the presence of both radial and nonradial pulsation modes. Note that nonradial and radial modes coexist in ν Eri and σ Sco. In these two cases Kubiak's (1980b) identification of the spherical harmonic degree of the dominant mode (see Table III) differs from Watson's (1988). In the case of β Cep, the A_{U-V}/A_V data are not inconsistent with Kubiak's (1980b) $l = 2$ identification. As can be seen from Figure 24, in this star, as well as in β Cru, α Vir,

τ^1 Lup, and λ Sco, the single observed mode is nonradial. The other singly-periodic variables plotted in this figure have the color-to-light amplitude ratios indicating $l = 0$. Finally, Figure 24 shows that an $l = 2$ mode is the dominant one in 12 Lac, in agreement with Smith's (1980) line-profile modelling (see Table III). For further applications of this method, see Heynderickx (1990, 1991).

The third method was worked out by Aerts *et al.* (1990) and pulsation parameters for δ Cet (radial mode), ν Eri (radial mode, and a $l = 2$, $m = -2$ second mode) and β Cru ($l = 3$, $m = -3$) were obtained.

These results support the last conclusion of the preceding section that in a multiperiodic β Cephei star different l modes can be simultaneously present. In addition, they show that the dominant mode in multiperiodic stars, or the only mode in singly-periodic ones, is not always radial.

An important case, discussed in some detail by Watson (1988) but not included in Figure 24, is that of V836 Cen. The light variation of this star, discovered by Waelkens and Rufener (1983a), consists of three short-period sinusoidal components. According to Watson (1988), all three components have color-to-light amplitude ratios favouring the $l = 1$ or $l = 2$ modes. In other words, V836 Cen is a multiperiodic β Cephei star in which no radial mode is seen.

5.3. The Period-Luminosity and Period-Luminosity-Color Relations

McNamara (1953) and, independently, Blaauw and Savedoff (1953) drew attention to the possible existence of a period-luminosity relation among the β Cephei stars. Blaauw and Savedoff (1953) based their suggestion on six stars, namely, δ Cet, β CMa, σ Sco, β Cep, 12 Lac, and 16 Lac, for which they derived distances from membership in associations or from the υ components of proper motion. Figure 25 is a reproduction of Blaauw and Savedoff's (1953) correlation, together with the result that one would obtain using M_V derived from β and c_0 by means of the calibration of Balona and Shobbrook (1984), presented in Section 4.3. As can be seen from the figure, the new absolute magnitudes lead to a considerably tighter correlation, but with a slope amounting to about half the original one.

A color term has been introduced into the relation by Leung (1967). Using 17 β Cephei stars with M_V known from spectrographic $H\gamma$ equivalent widths, he obtained the following, remarkably tight, relation:

$$\begin{array}{lllll} -5.72 \log P & + 10.4\,(B-V)_0 & + 2.60 & = M_V & \text{(P in hours)} \\ \pm 0.66 & \pm 3.1 & \pm 0.61 & \pm 0.17 & \end{array} \quad (11)$$

The period-luminosity and period-luminosity-color relations for the β Cephei stars have been subsequently derived by a number of workers, most recently by Jakate (1980) and Waelkens (1981).

The discovery of HR 3058 = QS Pup, HR 3088 = V372 Car, HR 3213 = IS Vel, which have periods much shorter than most other β Cephei stars but are nonetheless

TABLE III
The spherical harmonic mode identification from line-profile modeling

Multiperiod Variables						
Star	f (c/d)	K (km/s)	l	m	V_e (km/s)	i
σ Sco	4.0512	39.8	2	-1	70	50°
Kubiak (1980b)	4.1738	4.8	3	0		
12 Lac	4.2420	0.9	?	?	75	15°
Smith (1980)	5.0659	6.9	0	0		
	5.1791	18.5	2	0		
	5.3350	5.5	2	-1		
	5.4901	4.5	2	-2		
ν Eri	5.6202	9.2	1	+1	7	90°
Kubiak (1980b)	5.6375	5.4	1	0		
	5.6541	6.6	1	-1		
	5.7635	18.0	2	0		

Single-period Variables						
Star	f (c/d)	K (km/s)	l	m	V_e (km/s)	i
β Cep	5.2496	22.5	2	0	?	?
Kubiak (1978)						
BW Vul	4.9744	130	2	-1	130	60°
Kubiak (1978)						

located in the middle of the instability strip, led Jerzykiewicz and Sterken (1979, 1980) to reconsider the period-luminosity and period-luminosity-color relations. In Figure 26, reproduced from Jerzykiewicz and Sterken (1980), these three stars and NGC 6231-150 = HDE 326333 = V920 Sco, the shortest-period β Cephei variable known at the time, are shown as triangles. Including them in a least-squares solution would result in a period-luminosity relation with insignificant slope. Jerzykiewicz and Sterken (1980) concluded therefore that for β Cephei stars a universal period-luminosity relation does not exist.

Omitting the above-mentioned four stars yields the following period-luminosity relation:

$$\begin{array}{llll} -6.5 \log P & - \;10.8 & = & M_{bol} \\ \pm 1.8 & \pm\, 1.3 & & \pm 0.6 \end{array} \quad (12)$$

where P is the period, in days, of the component having the largest light amplitude. This relation is shown in Figure 26 by means of the straight line.

The large standard deviation on the right-hand side of the last equation can be considerably reduced by adding a $\log T_{\rm eff}$ term, analogous to the color term in eq. (11). Indeed, with the $\log T_{\rm eff}$ term included one gets:

$$\begin{array}{lllll} -3.4 \log P & - \;16.3 \log T_{\rm eff} & + \;62.9 & = & M_{bol} \\ \pm 0.9 & \pm\, 1.8 & \pm\, 8.1 & & \pm 0.27 \end{array} \quad (13)$$

Results of this kind are supposed to demonstrate that β Cephei variables obey a period-luminosity-color relation, just like the classical Cepheids do. However, as was pointed out by Jerzykiewicz and Sterken (1980), the decrease of the standard deviation between eqs. (12) and (13) is almost entirely due to the fact that for stars which were used to derive them there is already a tight $\log T_{\rm eff} - M_{bol}$ correlation:

$$\begin{array}{llll} -19.1 \log T_{\rm eff} & + \;77.6 & = & M_{bol} \\ \pm 2.1 & \pm\, 9.3 & & \pm 0.35 \end{array} \quad (14)$$

Therefore, eq. (13) should not be looked at as a considerable improvement over the period-luminosity relation, expressed by eq. (12), but as a barely significant modification of eq. (14), and consequently it does not prove that a period-luminosity-color relation exists, even for variables with $\log P > -0.9$.

5.4. The Pulsation Constant

Eliminating the mean radius of a pulsating star, $< R >$, between the definition of effective temperature and the period-mean density relation:

$$P \sqrt{< \rho >} = Q \quad (15)$$

where $< \rho >= 3\,{\rm M}/4\,\pi < R >^3$ is the mean density in solar units and Q is the pulsation constant (not to be confused with the photometric Q index introduced in Section 3.2), one gets:

$$M_{bol} = -10/3 \log P - 10 \log T_{\rm eff} - 5/3 \log {\rm M} + C' \quad (16)$$

where $C' = 10/3 \log Q + M_{bol\odot} + 10\, T_{eff\odot}$. If the mass, , is eliminated by means of the mass-luminosity relation

$$\log {\rm M} = m_1\, M_{bol} + m_0 \quad (17)$$

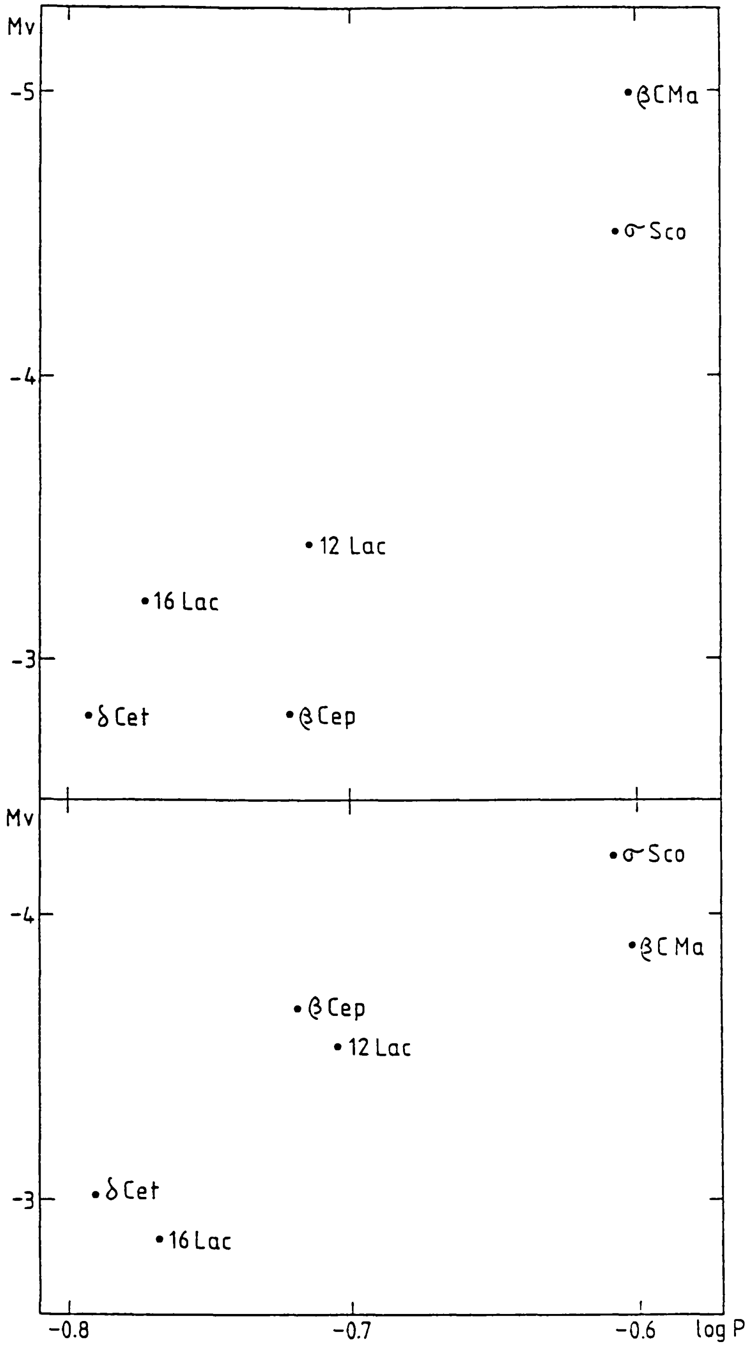

Fig. 25. Top: the period-luminosity correlation for the β Cephei stars according to Blaauw and Savedoff (1953). Bottom: the same correlation, plotted using modern absolute magnitudes.

eq. (16) becomes:

$$M_{bol} = -10/(3+5\,m_1)\,\log P - 30/(3+5\,m_1)\,\log T_{\rm eff} + C \tag{18}$$

where

$$C = (10\,\log Q + 3\,M_{bol\odot} + 30\,\log T_{eff\odot} - 5\,m_0)/(3+5\,m_1) \tag{19}$$

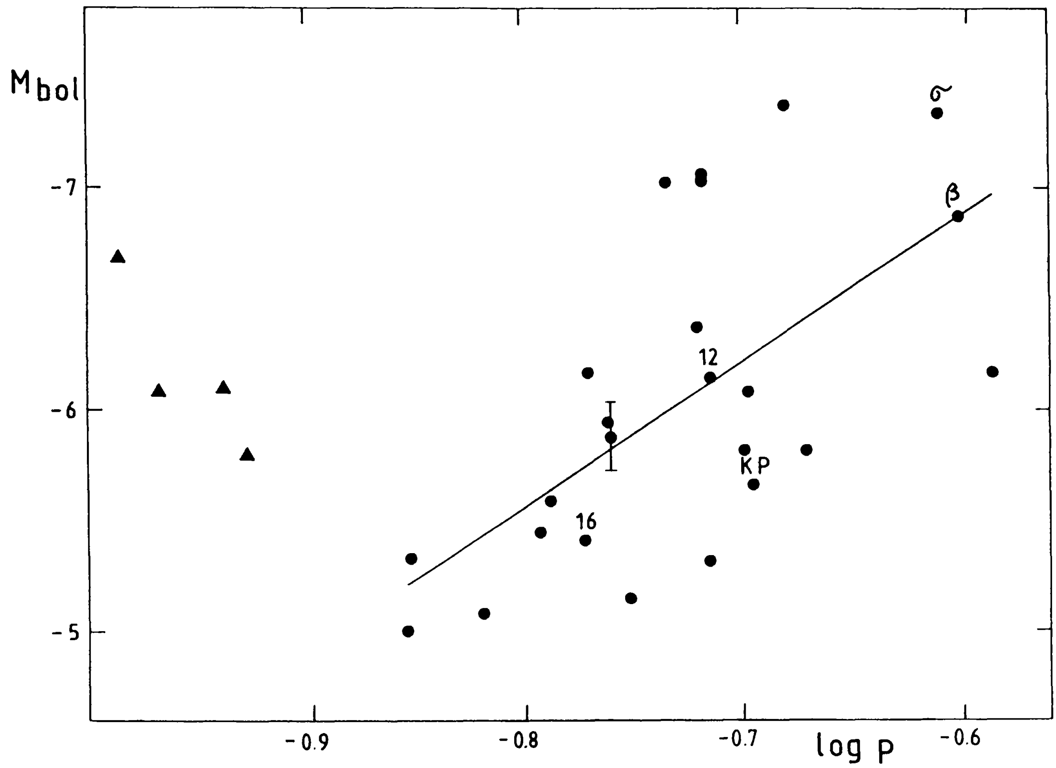

Fig. 26. The bolometric absolute magnitudes, M_{bol}, of the β Cephei stars, shown as a function of $\log P$, where P is the primary photometric period in days, and $M_{bol} = M_V + B.C.$ (from Jerzykiewicz and Sterken 1980). The M_V were obtained from β via the calibration of Crawford (1978), except that for the primary component of Spica (circle with the error bar) the absolute magnitude was obtained using the distance derived by Herbison-Evans *et al.* (1971) from the radial-velocity and interferometric observations of the system. Bolometric corrections were based on the scale of Code *et al.* (1976). The shortest-period variables, mentioned in the text, are plotted as triangles. The multiperiodic variables of Figure 22 are also indicated. The straight line corresponds to Eq. (11). Note that V986 Oph was omitted.

Assuming now that $\log T_{\rm eff} = t_1\, CI + t_0$ and $B.C. = b_1\, CI + b_0$, where CI is a color index, one can transform eq. (18) into the following $P - M_V - CI$ relation:

$$M_V = -10/(3 + 5\, m_1)\, \log P - [30\, t_1/(3 + 5\, m_1) + b_1]\, CI + D \qquad (20)$$

where $D = C - b_0 - 30\, t_0/(3 + 5\, m_1)$. Note that the coefficient at $\log P$ does not change between eqs. (18) and (20), and that the coefficients at $\log P$ and $\log Q$ are the same, except for the sign. In addition, they are functions of only the slope of the mass-luminosity relation.

The last equation can be used to derive an "observational" pulsation constant of any star for which the period of pulsation, visual absolute magnitude, and a suitable color index are known. Eq. (18) is used in the rare cases when the luminosity and effective temperature are known directly. Eq. (18) is also used when M_{bol} and $T_{\rm eff}$ are obtained from more complex calibration relations than linear functions of one color index.

For β Cephei stars, the observational pulsation constants were derived and compared with the theoretical ones, resulting from the pulsation calculations, by Hitotuyanagi and Takeuti (1964), Jones and Shobbrook (1974), Lesh and Aizenman (1974), Balona and Feast (1975), and others. Most recently, a thorough discussion of the pulsation constants of the β Cephei stars has been provided by Shobbrook (1985).

According to Hitotuyanagi and Takeuti (1964), the mean value of Q is equal to 0.021. Since the theoretical pulsation constants are equal to about 0.037, 0.028, and 0.022 for the radial fundamental (F), the radial first overtone ($1H$), and the radial second overtone ($2H$), respectively, this value would indicate the second overtone pulsation if the dominant pulsation mode were radial. However, Hitotuyanagi and Takeuti (1964) believed that the small value of Q can be better accounted for by a nonradial mode. Jones and Shobbrook (1974) deduced absolute magnitudes of several β Cephei stars from their membership in the Scorpio-Centaurus association and estimated absolute magnitudes of the remaining ones from the strength of the $H\gamma$ and $H\beta$ lines; they obtained pulsation constants indicating that two different modes must occur among these stars, probably F and $1H$. On the other hand, a recalibration of the $H\gamma$ absolute magnitudes and a discussion of observational errors led Balona and Feast (1975) to the conclusions that (1) the mean value of $Q = 0.027 \pm 0.001$, implying $1H$ pulsation, and (2) there is no need to postulate different modes of pulsation for the dominant oscillation in different stars. Lesh and Aizenman (1974), using a preliminary version of Crawford's (1978) calibration, derived pulsation constants indicating either $1H$ or $2H$ mode. Moreover, they noted that the pulsation constants are also consistent with the lowest-order nonradial acoustic modes, p_1 or p_2.

Most above-mentioned investigations were limited to less than 20 variables, mainly those that are referred to in the Introduction as the confirmed β Cephei stars. A significantly larger sample, including almost all presently known field β Cephei stars and such stars in NGC 6231-150, was used in plotting Figure 26 and deriving eqs. (12), (13), and (14). These stars are also represented in Figure 27, together with the theoretical $P - M - T_{\rm eff}$ relations. The latter were obtained from eq. (18), assuming the theoretical pulsation constants equal to 0.0372, 0.0278, and 0.022 for F, $1H$, and $2H$, respectively, and the mass-luminosity relation, eq. (17), with $m_1 = -0.135$ and $m_0 = 0.31$, derived from the hydrogen-burning phases of evolutionary tracks for the ridge line of the β Cephei strip in Figure 20.

The effective temperatures in Figure 27 were obtained from c_0 and the $c_0/T_{\rm eff}$ calibration of Davis and Shobbrook (1977), and the absolute magnitudes, from β and the β/M_V calibration of Crawford (1978). Spica and 16 Lac, the two stars mentioned in Section 4.2, can also be plotted in this figure without using c_0 and β. In the case of Spica, the empirical values of Code *et al.* (1976) place its primary component on the $1H$ line, while the photometric indices, corrected for the light of the secondary (Shobbrook 1978a), give $\log T_{\rm eff}$ and M_{bol} which indicate second-overtone pulsation (open circle and filled circle with error bars, respectively). If the

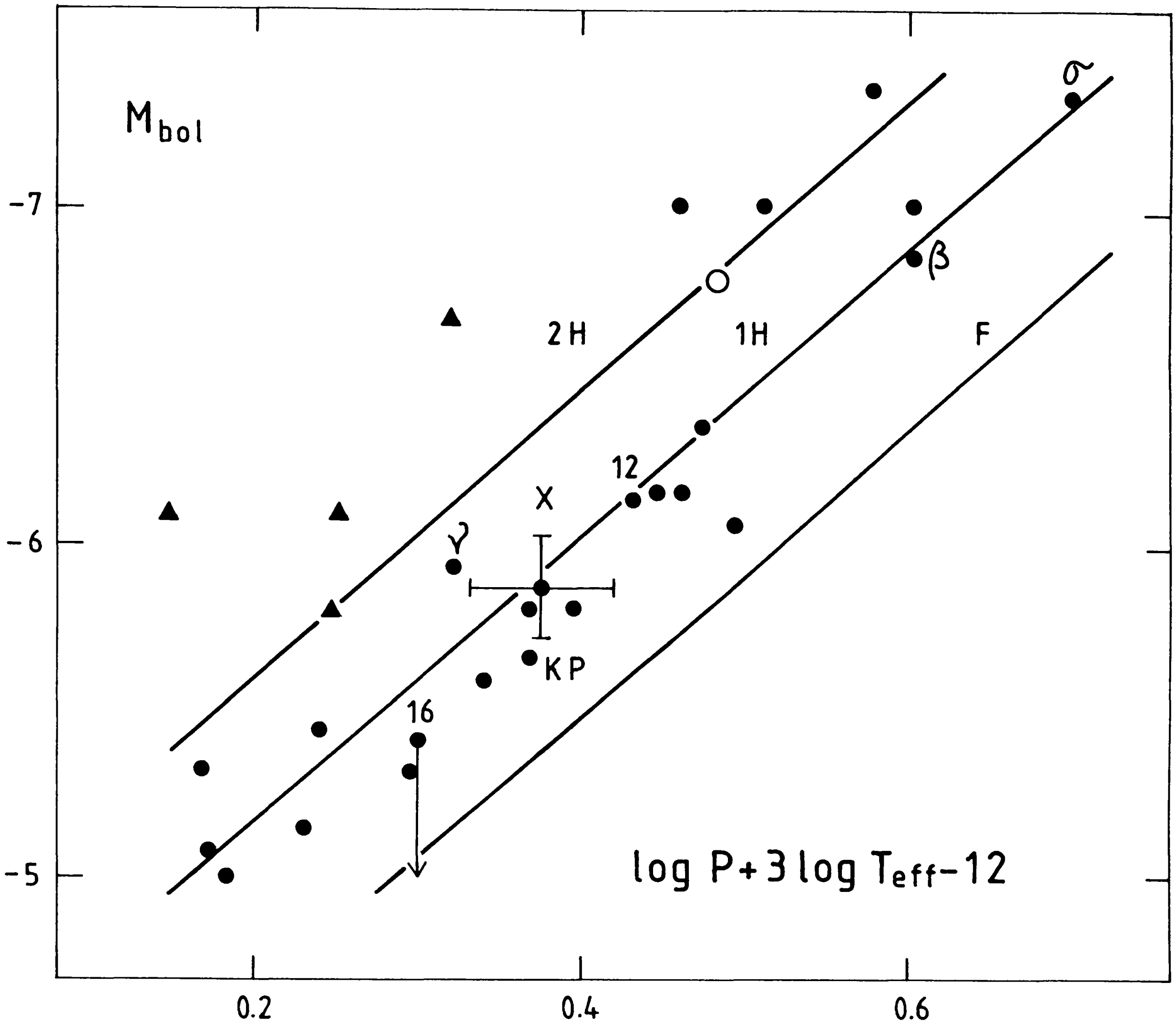

Fig. 27. The theoretical $P - M - T_{\rm eff}$ relations (straight lines) for the fundamental mode (F), the first overtone ($1H$), and the second overtone ($2H$), compared with observations (from Jerzykiewicz and Sterken 1980). The four shortest-period variables are plotted as triangles. The multiperiodic stars of Figs. 19 and 20 are indicated. See text for explanation of the remaining symbols.

interferometric radius and mass from Herbison-Evans *et al.* (1971) were used, the star would fall between the $2H$ and $1H$ lines, as indicated by a cross. In the case of 16 Lac, the radius and mass obtained from the eclipse solution (Jerzykiewicz 1980, Pigulski and Jerzykiewicz 1988) yield a pulsation constant close to 0.037. In Figure 27 this would move the star as shown by the arrow, provided that the photometric $T_{\rm eff}$ were correct.

In Section 4.2 we have already mentioned that these discrepancies would disappear if the absolute magnitudes based on Crawford's (1978) calibration were increased by about 0.5. Most points in Figure 27 would then fall close to the F line, but some, including the four shortest-period ones, would still indicate pulsation in the first or even second radial overtone. Using Shobbrook's (1983a) preliminary β/M_V calibration, given by eq. (8) of Section 4.2, would have the same effect. Shobbrook (1983a) has pointed out that if the lowest-Q variables were disregarded

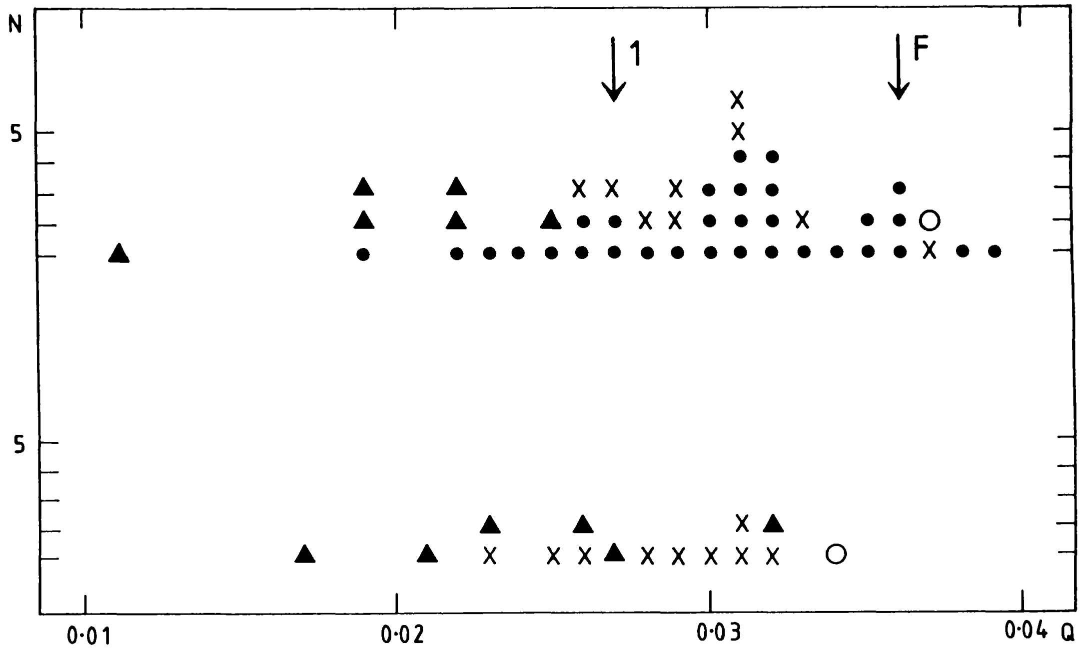

Fig. 28. Top: a histogram of the Q values based on the M_V calibration of Balona and Shobbrook (1984) and the $T_{\rm eff}$ and $B.C.$ calibrations of Balona (1984). Filled circles are field β Cephei stars, crosses are NGC 3293 variables, triangles are NGC 6231 variables, and the open circle is NGC 4755-F = BW Cru. Bottom: the same for the cluster variables using M_V from the clusters' mean distance moduli. The theoretical Q values for the fundamental and first overtone radial pulsation modes are indicated. After Shobbrook (1985).

on the grounds that their absolute magnitudes may be in error, the revision of the M_V scale would lead to the conclusion that the β Cephei stars are pulsating in the fundamental mode.

The last conclusion has been questioned by Shobbrook (1985) after the absolute-magnitude calibration of Balona and Shobbrook (1984), mentioned in Section 4.2, became available. Using this calibration, and Balona's (1984) $T_{\rm eff}$ and $B.C.$ scales (see Section 4.1), Shobbrook (1985) derived Q values for 47 β Cephei stars then known, including the NGC 3293, NGC 4755, and NGC 6231 variables. These Q values are shown in Figure 28 (upper histogram). In the same figure are also shown the Q values of the cluster variables, based on M_V obtained from the clusters' mean distance-moduli (lower histogram).

After analyzing the effect of observational errors on Q, Shobbrook (1985) concluded from the distributions shown in Figure 28 that the majority of β Cephei stars are pulsating in the first-overtone radial mode. He adds, however, that "possibly two stars (β Cen and 16 Lac) have fundamental-mode periods and several more have Q values much nearer that for the fundamental than for the first overtone", and notes that "the existence of higher-order radial, or nonradial, modes is likely in some cases even for the dominant period, particularly for HR 3213 = IS Vel."

6. The opacity mechanism and future observations

6.1. The Opacity Mechanism

As we mentioned in the Introduction, a quarter-century ago Christy (1966) pointed out that the region of ionization of He^+, responsible for destabilizing Cepheids and RR Lyrae variables, cannot be effective in the much hotter β Cephei stars. Christy (1966) made this remark in connection with Zhevakin's (1963) claim that pulsation calculations performed for Cepheid models can be extrapolated to β Cephei stars. Since then several instability mechanism were proposed, but none had been widely accepted. In particular, it was not known where in these stars the driving occurs. For example, Cox (1987) reviewed six deep-interior mechanisms and three envelope ones. He also suggested a new mechanism in which driving is provided by interaction between the outgoing flux and convection, periodically started in the sub-photospheric layers.

The sequence of events, leading to what is probably the correct solution of the problem, was the following. In a theoretical study of δ Scuti stars, Stellingwerf (1979) discovered a driving zone at temperatures near 150,000 K, generated by an opacity bump due to the He^+ ionization edge. He noted that although in the Cepheid instability strip this zone was too deep to have any effect on the overall pulsational driving, it could be effective in hotter stars. A detailed investigation (Stellingwerf 1978) showed that this "opacity bump" mechanism does tend to destabilize the radial fundamental mode in a region of the H-R diagram close to the observed position of β Cephei stars. However, actual instability was not found in models with the standard Los Alamos opacities. For the instability to appear, the opacity bump at $\log T = 5.2$ had to be arbitrarily enhanced. In view of the rather inadequate resolution of the opacity tables above 10^5 K, Stellingwerf (1978) argued that this modification may be reasonable. He also pointed out that the "opacity bump" driving does not depend on an ionization zone but on the changes of opacity with temperature and density; it is thus a pure κ - effect (Baker and Kippenhahn 1962). The instability strip obtained with the modified opacities runs approximately parallel to the observed one, but about 0.1 in $\log T_{\rm eff}$ toward the red. In addition, the overtones were strongly damped.

Stellingwerf's (1978) calculations were extended to nonradial modes by Saio and Cox (1980) and by Dziembowski and Kubiak (1981). For low $l > 0$ values these authors found the driving effect to be almost identical as for the radial modes of the same frequency. Dziembowski and Kubiak (1981) noted that although the effect is insufficient to cause instability in case when standard opacities are used, the assumption that it is actually responsible for the excitation provides a natural explanation of all major properties of β Cephei stars.

An important next step was taken by Simon (1982). In the first place, he showed that if the contribution of metals (that is, elements with atomic number greater than 2) to opacities in the temperature range between 10^5 and 10^6 K were increased by factors 2-3, classical Cepheid models would result which reproduce the observed ratios of the first- and second-overtone periods to that of the fundamental, thus removing the so called "mass discrepancy" for the double-mode and "bump" Cepheids. Then he noted that the run of opacity with temperature in case of increased metal contribution shows a bump similar to that postulated by Stellingwerf (1978) to energize β Cephei stars. Simon (1982) expressed his conclusion as follows:

> "It is thus quite possible that, by the single stroke of augmenting the heavy-element opacities by factors of 2-3, we can bring into line with the theory of stellar structure and evolution not only the double-mode and bump Cepheids, but the β Cephei pulsators as well."

Finally, he enumerated possible uncertainties in the Los Alamos opacity calculations and urged that the metal contribution to opacities be reexamined.

The reaction of the Los Alamos experts was swift and rather discouraging. Magee et al. (1984) flatly stated that as large an increase of the metal contribution to opacities as required by Simon (1982) is incompatible with the atomic physics. Slightly later, however, a very different message came from the Lawrence Livermore National Laboratory, where Iglesias *et al.* (1987) developed a new opacity code, which became known as OPAL. These authors concluded that an improved treatment of the atomic physics can significantly increase the opacities of metals in astrophysical mixtures. In subsequent papers Iglesias and his co-workers presented improved opacity tables for Cepheid models (Iglesias *et al.* 1990, Iglesias and Rogers 1991b), for the interior of the Sun (Iglesias and Rogers 1991a). A more extensive version of these tables was published by Rogers and Iglesias (1992). Very recently, Iglesias *et al.* (1992) showed that effects of spin-orbit interactions significantly enhance opacity in the region critical for driving pulsations of β Cephei stars.

The tables of Iglesias and Rogers (1991b) were used in an investigation of the Cepheid mass discrepancies by Moskalik *et al.* (1992). The result was that the new opacities lead to the double-mode Cepheid masses, in agreement with the stellar evolution theory, and that the discrepancy in the case of the bump Cepheids is greatly reduced.

In the context of the pulsation mechanism for β Cephei stars, a breakthrough occurred when Dziembowski and Pamyatnykh (1992a, 1992b) used the most recently updated OPAL tables of Iglesias et al. (1992). This work will be reviewed in the next section. The rest of the present section we devote to the results of Cox and Morgan (1990), Cox et al. (1991), Kiriakidis *et al.* (1992), and Moskalik and Dziembowski (1992), based on the earlier versions of the opacity tables.

Cox and his co-workers stressed the importance of "the sudden appearance of a tremendous number of same-shell transition iron lines as the temperature rises above 10^5 K." The model they found unstable had 12 $M_\odot$, 21000 $L_\odot$, 22750 K, $X = 0.70$, $Z = 0.02$, and arbitrarily-doubled iron abundance in the driving layers. On the evolutionary track in the theoretical H-R diagram this model is just before central hydrogen-exhaustion. Linear nonadiabatic calculations indicated instability

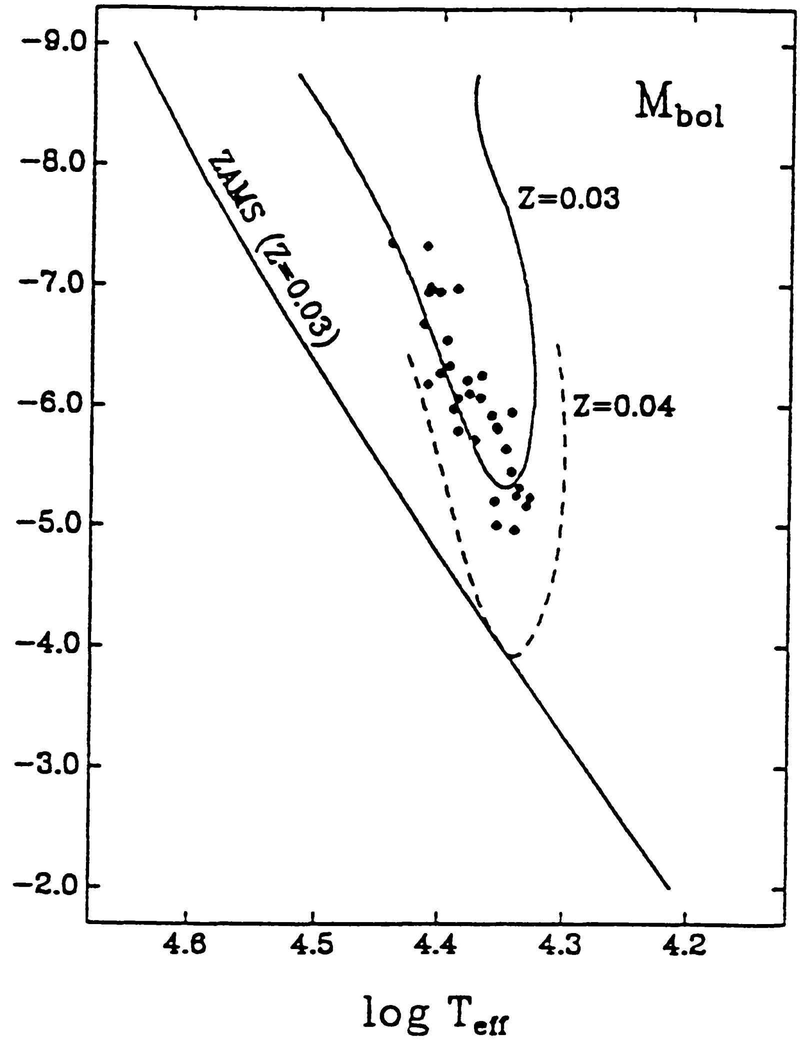

Fig. 29. The theoretical H-R diagram showing β Cephei stars (dots) and the domain of the radial fundamental mode instability for two values of the heavy-element content. From Moskalik and Dziembowski (1992).

of "the radial fundamental and one or two overtones as well as nonradial p-modes of low degree with periods (near 0.3 day) in the same range as the radial modes" (Cox and Morgan 1990) or "a radial mode and a few nonradial g-modes of low degree and order with periods near 0.3 day" (Cox *et al.* 1991). In both papers the authors conclude that "multimode behavior is theoretically expected for this pulsation mechanism, and for some B stars only a few nonradial modes (possibly selected by rotation) may survive to observable amplitudes." They also discuss, in a qualitative way, the connection between the iron abundance in the surface layers and such topics as nonvariable stars in the instability strip and the large pulsation-amplitude of BW Vul.

Kiriakidis *et al.* (1992) examined pulsational stability along 15 M$\odot$ evolutionary tracks with $Z = 0.02$ and 0.03, computed with the opacity tables of Rogers and Iglesias (1992). For $Z = 0.02$ they found all models to be stable, but for $Z = 0.03$

they obtained instability in the radial and nonradial fundamental p_1[6] modes up to $l = 4$ (they did not check higher l) in the interval $4.326 < \log T_{\rm eff} < 4.43$, that is, in the S-band region of the track (see Section 4.3). The overtones they found to be stable. Kiriakidis *et al.* (1992) remark that they were unable to confirm the Cox *et al.* (1991) results: all models along the 12 M$\odot$ track turned out to be stable.

The opacity data of Rogers and Iglesias (1992) were also used by Moskalik and Dziembowski (1992), who conducted an extensive pulsation-stability survey of envelope models in the parameter range corresponding to β Cephei stars. Their conclusions may be summarized as follows:

1. The fundamental radial mode becomes unstable for $Z > 0.03$.
2. The width of the instability strip depends on the heavy-element content; this is illustrated in Figure 29, which shows that an instability strip for Z somewhat higher than 0.03 would best fit the points representing β Cephei stars.
3. The mechanism favours low-frequency oscillations: only modes with $Q > 0.032$ may be unstable.
4. The theoretical instability strip extends above the β Cephei region in the H-R diagram. Thus, the same driving mechanism may be at work in the Luminous Blue Variables[7].

6.2. THE BREAKTHROUGH

In their I.A.U. Colloquium 137 poster, Dziembowski and Pamyatnykh (1992a) report: "In a very recent work Iglesias, Rogers and Wilson (1992) showed that effects of spin-orbit interactions significantly enhance opacity in the region critical for driving pulsations in β Cephei stars. These effects and improved information about the solar metal-mixture have been included in the updated opacity tables. The tables have been kindly provided to us via electronic mail by Dr. Rogers. Only thanks to the use of this efficient way of communication we managed to remake our calculations before this conference."

Dziembowski and Pamyatnykh (1992a, 1992b) show that the consequences for β Cephei models of this update of the OPAL opacities are important indeed. Instability is now obtained also for $Z = 0.02$. In fact, most β Cephei stars lie in the region of the H-R diagram where models with $Z = 0.02$ are pulsationally unstable because of the κ -mechanism due to the metal opacity-bump at temperatures about 2×10^5 K. Only for variables in NGC 6231 a higher metallicity may be required. As in the case of earlier calculations of Moskalik and Dziembowski (1992), instability persists beyond the end of core hydrogen-burning.

Instability is no longer confined to the fundamental mode, as was the case for the older calculations mentioned in the preceding section, but overtones are also predicted to be unstable. This very important result is illustrated in Figs. 30 and

[6] In this section we use the recently adopted notation in which the radial modes F, $1H$, $2H$, etc. are denoted p_1, p_2, p_3, etc., and the nonradial f mode is denoted p_0 (see below, Figs. 30 and 31).

[7] for observational evidence supporting this hypothesis, see Sterken (1988b).

31 where the dimensionless frequency $\sigma = 0.06692/Q$ is plotted as a function of $\log T_{\rm eff}$.

As can be seen from Figure 30, a separate instability region appears for $l = 6$ and $l = 8$ at $\sigma < 1$ and $\log T_{\rm eff} < 4.37$. The unstable modes in this region are g-modes of rather high order. Dziembowski and Pamyatnykh (1992b) note that these long-period modes of high l may account for the line-profile variability discovered recently in several Be stars by Vogt and Penrod (1983), Baade (1984), and others.

As far as β Cephei stars are concerned, Dziembowski and Pamyatnykh (1992b) express their conclusion as follows:

> "There is little doubt that the classical κ -mechanism acting in the zone of the bump in metal opacity may account for excitation of all modes discovered in β Cephei stars. The problem we are now facing is to explain why only some of the unstable modes reach observable amplitudes. Thus, our understanding of this type of variable stars has reached a similar level as that of the δ Scuti stars. There is also a large similarity in the prospects for testing theory of stellar evolution connected with studies of these two types of objects."

6.3. Future Observations

The connection between β Cephei-type pulsations and stellar metallicity should be investigated observationally. In fact, Simon's (1982) hypothesis and the theoretical work that followed (see Section 6.1) has stimulated the photometric searches for β Cephei stars in the Magellanic Clouds by Sterken and Jerzykiewicz (1988) and Balona (1992a, 1992b). The negative results, especially those of the extensive observations in the SMC cluster NGC 330 (Balona 1992a) and in the LMC clusters NGC 2004 and NGC 2100 (Balona 1992b), provide a strong – albeit indirect – argument in favour of the metal opacity-driving in β Cephei stars.

A negative result was also obtained by Kubiak (1990) in his β Cephei search in the LMC cluster NGC 1712. Kubiak (1990) accounts for the lack of β Cephei stars in this cluster by its very young age.

As far as the galactic β Cephei stars are concerned, Waelkens *et al.* (1991) have suggested that the locus of the instability strip depends on metallicity in the sense that at higher metallicity the strip is displaced toward the blue. They based this suggestion on three premises:

1. there is an increase of the average metallicity with decreasing galactocentric distance (Mayor 1976, Janes 1979).
2. all β Cephei stars discovered so far in the Sagittarius spiral arm, including members of NGC 6231, are among the hottest objects of this type.
3. their search for cool β Cephei variables in the Sagittarius arm was unsuccessful.

Future work along these lines is certainly needed. An attempt should be made to derive observationally the minimum value of Z at which β Cephei pulsations still occur. The result could be used as a test of the opacity calculations. However, metal abundance should be inferred more directly than from the star's galactic position, because the latter may be misleading. An example is provided by PHL 346, a

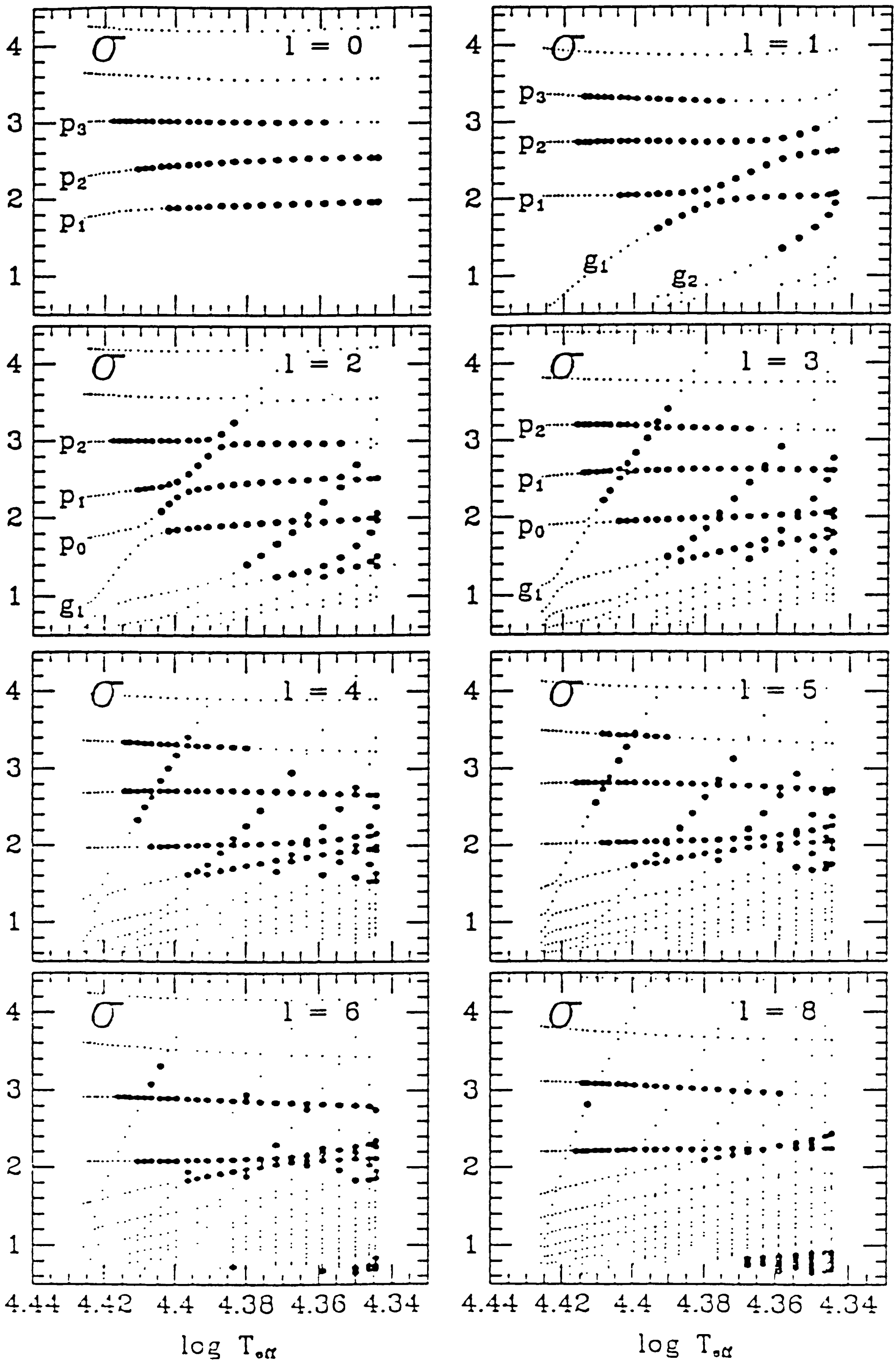

Fig. 30. Dimensionless frequencies, $\sigma = 0.06692/Q$, of low-order modes for models of a 12 M$\odot$, $Z = 0.03$ star in the main-sequence evolutionary phase. The small and big dots correspond to stable and unstable modes, respectively. From Dziembowski and Pamyatnykh (1992b).

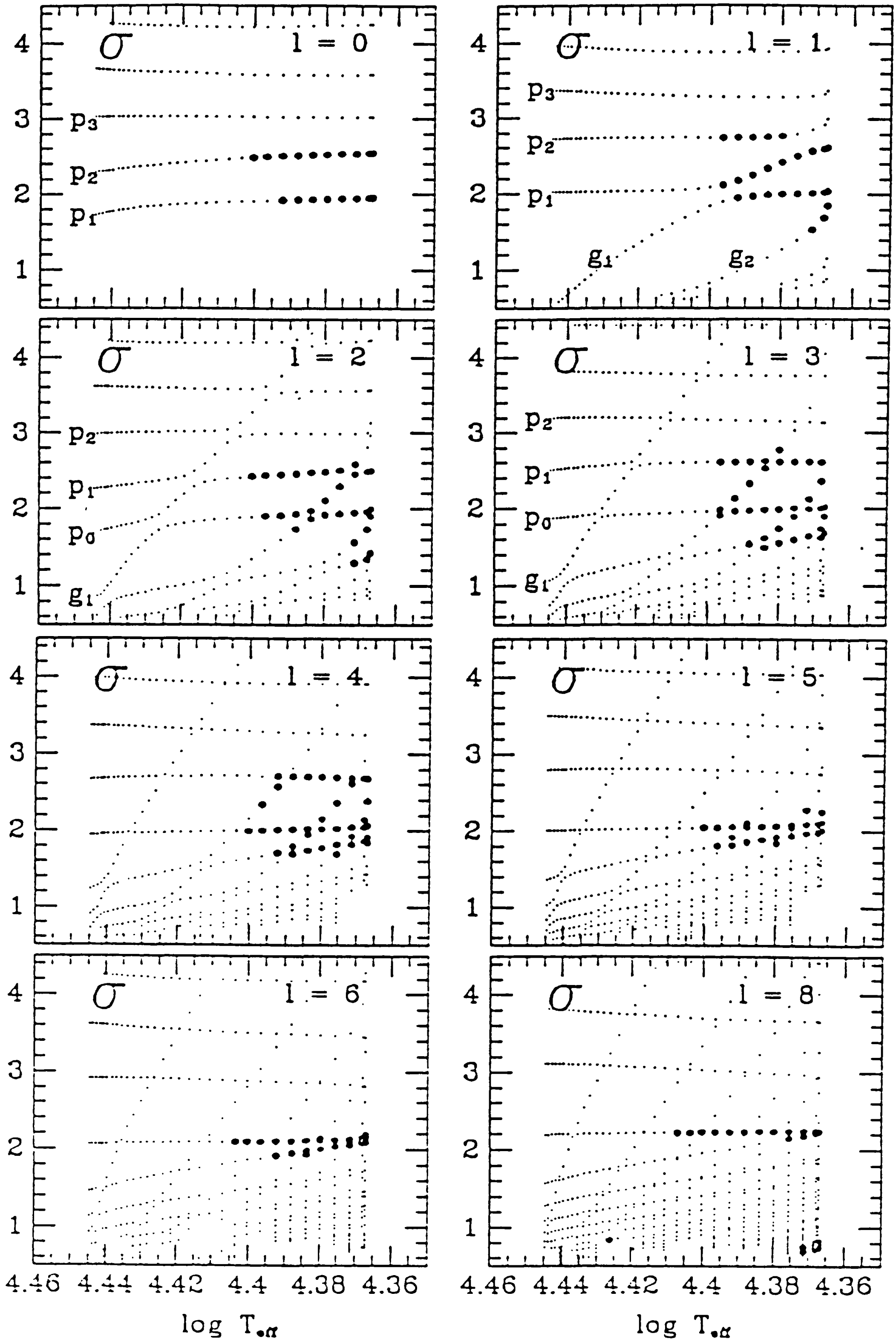

Fig. 31. The same as Figure 30, but for $Z = 0.02$. From Dziembowski and Pamyatnykh (1992b).

β Cephei variable situated at a distance of about -9 kpc from the galactic plane (Waelkens and Rufener 1988, Kilkenny and van Wyk 1990), which was shown by Keenan at al. (1986) to have essentially solar abundance of heavy elements.

Observations may help to answer the question whether pulsation amplitudes of β Cephei stars are determined by the iron abundance in their surface layers, as proposed by Cox and Morgan (1990), or by resonant mode-coupling, as suggested by Dziembowski (1988). One approach would involve an examination of the amplitude versus metallicity connection. For example, the fact that the NGC 6231 variables have very small amplitudes could be used as an argument against the Cox and Morgan (1990) hypothesis if it were shown that (1) these stars are indeed large Z objects, and (2) they are low l pulsators. Another possibility would consist in investigating extreme cases, that is, constant stars in the instability strip on one hand, and the large-amplitude variables such as BW Vul, on the other.

Finally, there are the prospects of testing observationally the theory of stellar evolution, mentioned in the quotation from Dziembowski and Pamyatnykh (1992b) at the end of the preceding section. These prospects are connected with the prediction, seen in Figs. 30 and 31, that the instability extends to g-modes, and the fact that frequencies of such modes are sensitive to the treatment of the convective-core boundary. Thus, another important task for observers is the detection of low order g-modes in β Cephei stars.

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Appendix

In the following tables are listed the photometric indices (Tables IA to IIIA) and the effective temperatures and absolute bolometric magnitudes (Table IVA) of the β Cephei stars of Table I.

TABLE I(A)
UBV color indices and the *Q* values of β Cephei stars.

HD/DM	Name	MK	$B-V$	$U-B$	Q	Source
886	γ Peg	B2 IV	−0.229	−0.862	−0.709	(1), (2), (3)
16582	δ Cet	B2 IV	−0.215	−0.864	−0.720	(1), (4)
21803	KP Per	B2 IV	0.032	−0.711	−0.733	(1), (2), (5)
29248	ν Eri	B2 III	−0.210	−0.883	−0.742	(1), (2)
44743	β CMa	B1 II−III	−0.24	−0.97	−0.807	(1)
46328	ξ^1 CMa	B1 III	−0.24	−0.98	−0.817	(1)
50707	15 EY CMa	B1 III	−0.215	−0.95	−0.804	(1)
52918	19 V637 Mon	B1 IV	−0.200	−0.920	−0.784	(1), (2)
56014	27 EW CMa	B3 IIIp	−0.20	−0.69	−0.557	(1)
59864	—	B1 III	−0.09	−0.84	−0.778	(6)
61068	PT Pup	B2 III	−0.19	−0.90	−0.771	(6)
63949	QS Pup	B1.5 IV	−0.14	−0.86	−0.765	(6)
64365	QU Pup	B2 IV	−0.20	−0.85	−0.716	(6)
64722	V372 Car	B1.5 IV	−0.15	−0.90	−0.798	(6)
68324	IS Vel	B2 IVn	−0.21	−0.90	−0.758	(6)
78616	KK Vel	B2 II−III	−0.02	−0.75	−0.736	(6)
80383	IL Vel	B2 III	—	—	—	
303068	V400 Car	B1 V	0.005	−0.810	−0.813	(9), (10)
303067	V401 Car	B0.5 V	0.035	−0.755	−0.779	(9), (10)
−57 3500	V403 Car	B0.5 III−V	−0.005	−0.810	−0.807	(9), (10)
—	V412 Car	—	0.03	−0.75	−0.771	(9)
−57 3506	V404 Car	B1 III	0.025	−0.785	−0.802	(9), (10)
−57 3507	V405 Car	B1 IV	−0.045	−0.855	−0.824	(9), (10)
−57 3517	V378 Car	B1 III	0.050	−0.790	−0.825	(9), (10)
−57 3524	V406 Car	B1 IV	−0.040	−0.805	−0.778	(9), (10)
92007	V380 Car	B0.5 III	0.060	−0.735	−0.777	(9), (10)
92024	V381 Car	B1 III	−0.030	−0.865	−0.844	(9), (10)
111123	β Cru	B0.5 III	−0.23	−0.98	−0.823	(1)
−59 4564	BW Cru	B2 III	0.09	−0.70	−0.762	(11)
112481	V856 Cen	B2 II−III	0.04	−0.74	−0.768	(15)
116658	α Vir	B1 IV	−0.23	−0.93	−0.775	(1)
118716	ϵ Cen	B1 III	−0.22	−0.92	−0.771	(1)
122451	β Cen	B1 III	−0.22	−0.98	−0.830	(7)
126341	τ^1 Lup	B2 IV	−0.15	−0.78	−0.680	(1)
129056	α Lup	B1.5 III	−0.20	−0.88	−0.745	(1)
129557	BU Cir	B2 III	−0.06	—	—	(1)
129929	V836 Cen	B3 V	−0.18	−0.87	−0.748	(15)
136298	δ Lup	B1.5 IV	−0.22	−0.88	−0.732	(1)
145794	V349 Nor	B2 II−III	—	—	—	
147165	σ Sco	B1 III	0.135	−0.695	−0.789	(1)
147985	V348 Nor	B1−2 II−III	—	—	—	
−41 7706	V945 Sco	B1:V + B1:V	0.17	−0.63	−0.748	(12)
−41 7711	—	B2 V + B2 V	0.174	−0.607	−0.728	(14)
−41 7715	V946 Sco	B2 IV−Vn	0.12	−0.47	−0.550	(12)

Table I(A) (continued)

HD/DM	Name	MK	$B-V$	$U-B$	Q	Source
326330	V964 Sco	B1 Vn	0.178	−0.674	−0.799	(14)
−41 7753	V947 Sco	B1 V	0.25	−0.63	−0.807	(13)
326333	V920 Sco	B1 Vn	0.20	−0.62	−0.760	(13)
156662	V831 Ara	B2 III	0.17	−0.65	−0.769	(6)
157056	θ Oph	B2 IV	−0.215	−0.845	−0.702	(1)
158926	λ Sco	B1.5 IV	−0.22	−0.89	−0.742	(1)
160578	κ Sco	B1.5 III	−0.22	−0.885	−0.737	(1)
163472	V2052 Oph	B2 IV−V	0.09	−0.65	−0.712	(2)
165174	V986 Oph	B0 IIIn	−0.009	−0.909	−0.903	(1), (2), (8)
166540	—	B0.5 IV	0.16	−0.73	−0.843	(16)
199140	BW Vul	B2 III	−0.15	−0.90	−0.798	(2)
205021	β Cep	B1 III	−0.225	−0.960	−0.807	(1), (2)
214993	12 DD Lac	B1.5 III	−0.140	−0.87	−0.775	(1), (2)
—	PHL 346	—	—	—	—	
216916	16 EN Lac	B2 IV	−0.145	−0.835	−0.737	(1), (2)

(1) Johnson *et al.* (1966)
(2) Crawford *et al.* (1971b)
(3) Jerzykiewicz (1970)
(4) Jerzykiewicz (1971a)
(5) Jerzykiewicz (1971b)
(6) Deutschman *et al.* (1976)
(7) Hogg (1958)
(8) Jerzykiewicz (1975)
(9) Turner *et al.* (1980)
(10) Feinstein and Marraco (1980)
(11) Perry *et al.* (1976)
(12) Feinstein and Ferrer (1968)
(13) Schild *et al.* (1969)
(14) Garrison and Schild (1979)
(15) Hill (1970)
(16) Guetter (1974)

TABLE II(A)

The c_0 and β indices of β Cephei stars

HR/Cluster	HD/DM	Name	MK	c_0	β
39	886	γ Peg	B2 IV	0.117	2.626
779	16582	δ Cet	B2 IV	0.092	2.619
1072	21803	KP Per	B2 IV	0.076	2.615
1463	29248	ν Eri	B2 III	0.063	2.610
2294	44743	β CMa	B1 II–III	−0.008	2.593
2387	46328	ξ^1 CMa	B1 III	−0.028	2.588
2571	50707	15 EY CMa	B1 III	−0.022	2.594
2648	52918	19 V637 Mon	B1 IV	0.012	2.592
2745	56014	27 EW CMa	B3 IIIp	0.157	2.569
—	59864	—	B1 III	−0.007	2.600
2928	61068	PT Pup	B2 III	0.038	2.619
3058	63949	QS Pup	B1.5 IV	0.014	2.616
3078	64365	QU Pup	B2 IV	0.104	2.621
3088	64722	V372 Car	B1.5 IV	0.007	2.610
3213	68324	IS Vel	B2 IVn	0.046	2.608
—	78616	KK Vel	B2 II–III	—	—
—	80383	IL Vel	B2 III	0.060	2.626
N3293–11	303068	V400 Car	B1 V	0.000	2.611
N3293–10	303067	V401 Car	B0.5 V	−0.008	2.604
N3293–16	−57 3500	V403 Car	B0.5 III–V	−0.022	2.592
N3293–65	—	V412 Car	—	0.020	2.585:
N3293–23	−57 3506	V404 Car	B1 III	−0.022	2.604
N3293–14	−57 3507	V405 Car	B1 IV	−0.024	2.596
N3293–24	−57 3517	V378 Car	B1 III	−0.051	2.594
N3293–18	−57 3524	V406 Car	B1 IV	0.001	2.606
N3293–27	92007	V380 Car	B0.5 III	−0.004	2.600
N3293–5	92024	V381 Car	B1 III	−0.035	2.600
4853	111123	β Cru	B0.5 III	−0.045	2.597
N4755–F	−59 4564	BW Cru	B2 III	0.000	2.606
—	112481	V856 Cen	B2 II–III	—	—
5056	116658	α Vir	B1 IV	0.018	2.607
5132	118716	ϵ Cen	B1 III	0.038	2.610
5267	122451	β Cen	B1 III	−0.010	2.594
5395	126341	τ^1 Lup	B2 IV	0.122	2.622
5469	129056	α Lup	B1.5 III	0.093	2.608
5488	129557	BU Cir	B2 III	0.022	2.616
—	129929	V836 Cen	B3 V	—	—
5695	136298	δ Lup	B1.5 IV	0.082	2.616
—	145794	V349 Nor	B2 II–III	—	—
6084	147165	σ Sco	B1 III	−0.052	2.590
—	147985	V348 Nor	B1–2 II–III	—	—
N6231–253	−41 7706	V945 Sco	B1:V + B1:V	−0.046	2.598
N6231–282	−41 7711	—	B2 V + B2 V	0.050	2.622
N6231–261	−41 7715	V946 Sco	B2 IV–Vn	0.063	2.626

Table II(A) (continued)

HR/Cluster	HD/DM	Name	MK	c_0	β
N6231–238	326330	V964 Sco	B1 Vn	−0.062	2.615
N6231–110	−41 7753	V947 Sco	B1 V	−0.072	2.612
N6231–150	326333	V920 Sco	B1 Vn	−0.052	2.607
—	156662	V831 Ara	B2 III	—	2.614
6453	157056	θ Oph	B2 IV	0.100	2.621
6527	158926	λ Sco	B1.5 IV	0.073	2.612
6580	160578	κ Sco	B1.5 III	0.071	2.612
6684	163472	V2052 Oph	B2 IV–V	0.089	2.630
6747	165174	V986 Oph	B0 IIIn	−0.170	2.565
—	166540	—	B0.5 IV	—	—
8007	199140	BW Vul	B2 III	0.006	2.611
8238	205021	β Cep	B1 III	0.004	2.605
8640	214993	12 DD Lac	B1.5 III	0.028	2.608
—	—	PHL 346	—	—	—
8725	216916	16 EN Lac	B2 IV	0.070	2.621

TABLE III(A)
The photometric X, Y, Z coordinates of β Cephei stars derived from Geneva photometry (Rufener 1988).

HR/Cluster	HD/DM	Name	MK	X	Y	Z
39	886	γ Peg	B2 IV	0.362	0.025	−0.003
779	16582	δ Cet	B2 IV	0.352	0.006	0.006
1072	21803	KP Per	B2 IV	0.302	0.024	−0.013
1463	29248	ν Eri	B2 III	0.304	0.016	0.007
2294	44743	β CMa	B1 II–III	0.190	0.031	0.000
2387	46328	ξ^1 CMa	B1 III	0.178	0.023	0.013
2571	50707	15 EY CMa	B1 III	0.202	0.019	0.010
2648	52918	19 V637 Mon	B1 IV	0.238	0.019	0.012
2745	56014	27 EW CMa	B3 IIIp	0.440	0.064	0.007
—	59864	—	B1 III	—	—	—
2928	61068	PT Pup	B2 III	0.276	0.005	0.006
3058	63949	QS Pup	B1.5 IV	0.248	−0.003	0.002
3078	64365	QU Pup	B2 IV	0.354	0.009	0.004
3088	64722	V372 Car	B1.5 IV	0.234	−0.001	0.000
3213	68324	IS Vel	B2 IVn	0.274	0.012	0.002
—	78616	KK Vel	B2 II–III	0.280	0.002	−0.007
—	80383	IL Vel	B2 III	0.318	0.015	0.002
N3293–11	303068	V400 Car	B1 V	—	—	—
N3293–10	303067	V401 Car	B0.5 V	0.206	0.001	−0.009
N3293–16	−57 3500	V403 Car	B0.5 III–V	—	—	—
N3293–65	—	V412 Car	—	0.219	−0.032	0.003
N3293–23	−57 3506	V404 Car	B1 III	0.184	−0.013	−0.001
N3293–14	−57 3507	V405 Car	B1 IV	—	—	—
N3293–24	−57 3517	V378 Car	B1 III	—	—	—
N3293–18	−57 3524	V406 Car	B1 IV	—	—	—
N3293–27	92007	V380 Car	B0.5 III	—	—	—
N3293–5	92024	V381 Car	B1 III	—	—	—
4853	111123	β Cru	B0.5 III	0.164	0.027	0.003
N4755–F	−59 4564	BW Cru	B2 III	0.232	0.027	−0.003
—	112481	V856 Cen	B2 II–III	0.180	0.005	0.012
5056	116658	α Vir	B1 IV	0.255	0.012	0.010
5132	118716	ϵ Cen	B1 III	0.250	0.015	0.005
5267	122451	β Cen	B1 III	0.198	0.023	0.003
5395	126341	τ^1 Lup	B2 IV	0.366	0.009	0.002
5469	129056	α Lup	B1.5 III	0.328	0.022	0.007
5488	129557	BU Cir	B2 III	0.236	0.000	−0.003
—	129929	V836 Cen	B3 V	0.272	0.004	0.002
5695	136298	δ Lup	B1.5 IV	0.296	0.029	0.008
—	145794	V349 Nor	B2 II–III	0.184	0.029	−0.006
6084	147165	σ Sco	B1 III	0.162	0.054	−0.007
—	147985	V348 Nor	B1–2 II–III	0.258	0.011	−0.015
N6231–253	−41 7706	V945 Sco	B1:V + B1:V	0.179	0.009	−0.011
N6231–282	−41 7711	—	B2 V + B2 V	0.283	0.025	−0.017
N6231–261	−41 7715	V946 Sco	B2 IV–Vn	0.265	−0.012	−0.025

Table III(A) (continued)

HR/Cluster	HD/DM	Name	MK	*X*	*Y*	*Z*
N6231–238	326330	V964 Sco	B1 Vn	0.134	−0.008	−0.017
N6231–110	−41 7753	V947 Sco	B1 V	0.150	0.019	−0.011
N6231–150	326333	V920 Sco	B1 Vn	0.166	0.010	−0.001
—	156662	V831 Ara	B2 III	0.190	0.034	−0.010
6453	157056	θ Oph	B2 IV	0.342	0.014	0.004
6527	158926	λ Sco	B1.5 IV	0.290	0.014	0.002
6580	160578	κ Sco	B1.5 III	0.288	0.032	0.003
6684	163472	V2052 Oph	B2 IV–V	0.340	0.011	−0.001
6747	165174	V986 Oph	B0 IIIn	0.002	0.030	−0.001
—	166540	—	B0.5 IV	0.106	0.009	−0.005
8007	199140	BW Vul	B2 III	0.262	0.031	0.001
8238	205021	β Cep	B1 III	0.207	0.030	0.003
8640	214993	12 DD Lac	B1.5 III	0.260	0.029	0.000
—	—	PHL 346	—	0.319	0.008	0.006
8725	216916	16 EN Lac	B2 IV	0.315	0.016	0.001

TABLE IV(A)

Log $T_{\rm eff}$ and M_{bol} of β Cephei stars derived from c_0 and β by means of the formulae of Balona (1984) and Balona and Shobbrook (1984).

HR/Cluster	HD/DM	Name	MK	log $T_{\rm eff}$	M_{bol}
39	886	γ Peg	B2 IV	4.305	−4.57
779	16582	δ Cet	B2 IV	4.317	−4.85
1072	21803	KP Per	B2 IV	4.325	−5.03
1463	29248	ν Eri	B2 III	4.330	−5.22
2294	44743	β CMa	B1 II–III	4.371	−6.13
2387	46328	ξ^1 CMa	B1 III	4.384	−6.44
2571	50707	15 EY CM	B1 III	4.385	−6.23
2648	52918	19 V637 Mon	B1 IV	4.353	−6.00
2745	56014	27 EW CMa	B3 IIIp	—	—
—	59864	—	B1 III	4.377	−5.96
2928	61068	PT Pup	B2 III	4.354	−5.20
3058	63949	QS Pup	B1.5 IV	4.371	−5.44
3078	64365	QU Pup	B2 IV	4.310	−4.74
3088	64722	V372 Car	B1.5 IV	4.373	−5.62
3213	68324	IS Vel	B2 IVn	4.341	−5.37
—	78616	KK Vel	B2 II–III	—	—
—	80383	IL Vel	B2 III	4.342	−4.90
N3293–11	303068	V400 Car	B1 V	4.380	−5.66
N3293–10	303067	V401 Car	B0.5 V	4.381	−5.88
N3293–16	−57 3500	V403 Car	B0.5 III–V	4.383	−6.28
N3293–65	—	V412 Car	—	4.337:	−6.15:
N3293–23	−57 3506	V404 Car	B1 III	4.394	−6.01
N3293–14	−57 3507	V405 Car	B1 IV	4.389	−6.21
N3293–24	−57 3517	V378 Car	B1 III	4.414	−6.53
N3293–18	−57 3524	V406 Car	B1 IV	4.375	−5.76
N3293–27	92007	V380 Car	B0.5 III	4.374	−5.93
N3293–5	92024	V381 Car	B1 III	4.403	−6.22
4853	111123	β Cru	B0.5 III	4.411	−6.40
N4755–F	−59 4564	BW Cru	B2 III	4.376	−5.77
—	112481	V856 Cen	B2 II–III	—	—
5056	116658	α Vir	B1 IV	4.362	−5.60
5132	118716	ϵ Cen	B1 III	4.348	−5.38
5267	122451	β Cen	B1 III	4.374	−6.13
5395	126341	τ^1 Lup	B2 IV	4.300	−4.63
5469	129056	α Lup	B1.5 III	4.309	−5.10
5488	129557	BU Cir	B2 III	4.365	−5.38
—	129929	V836 Cen	B3 V	—	—
5695	136298	δ Lup	B1.5 IV	4.321	−4.97
—	145794	V349 Nor	B2 II–III	—	—
6084	147165	σ Sco	B1 III	4.411	−6.63
—	147985	V348 Nor	B1–2 II–III	—	—
N6231–253	−41 7706	V945 Sco	B1:V + B1:V	4.413	−6.39
N6231–282	−41 7711	—	B2 V + B2 V	4.347	−5.05
N6231–261	−41 7715	V946 Sco	B2 IV–Vn	4.340	−4.88

Table IV(A) (continued)

HR/Cluster	HD/DM	Name	MK	log T_{eff}	M_{bol}
N6231–238	326330	V964 Sco	B1 Vn	4.442	−6.24
N6231–110	−41 7753	V947 Sco	B1 V	4.452	−6.42
N6231–150	326333	V920 Sco	B1 Vn	4.426	−6.27
—	156662	V831 Ara	B2 III	—	—
6453	157056	θ Oph	B2 IV	4.313	−4.76
6527	158926	λ Sco	B1.5 IV	4.325	−5.11
6580	160578	κ Sco	B1.5 III	4.326	−5.12
6684	163472	V2052 Oph	B2 IV–V	4.324	−4.64
6747	165174	V986 Oph	B0 IIIn	4.574	−9.31
—	166540	—	B0.5 IV	—	—
8007	199140	BW Vul	B2 III	4.375	−5.61
8238	205021	β Cep	B1 III	4.372	−5.75
8640	214993	12 DD Lac	B1.5 III	4.355	−5.50
—	—	PHL 346	—	—	—
8725	216916	16 EN Lac	B2 IV	4.333	−4.94