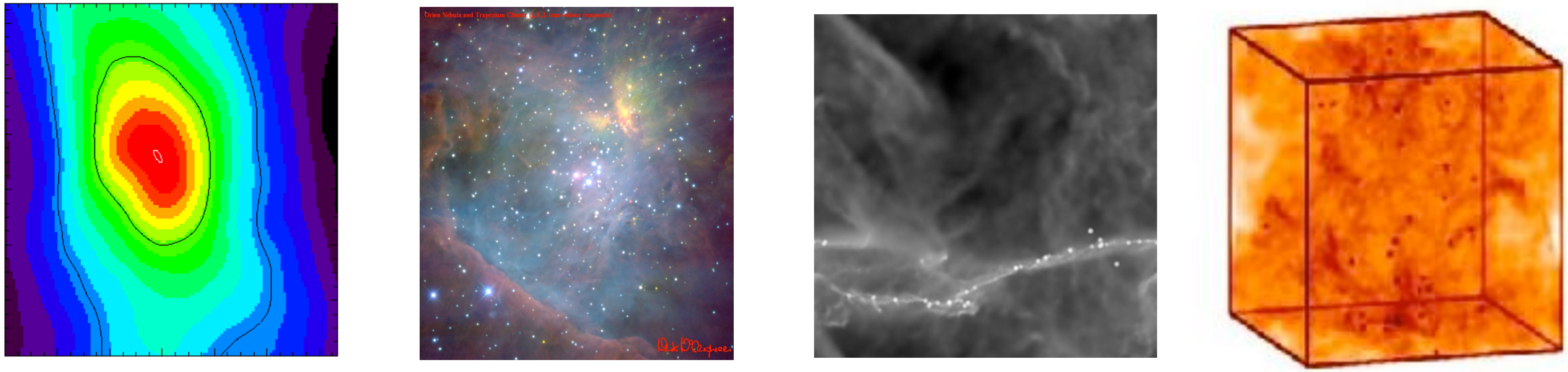
Sternentstehung



Ralf Klessen

Zentrum für Astronomie Heidelberg

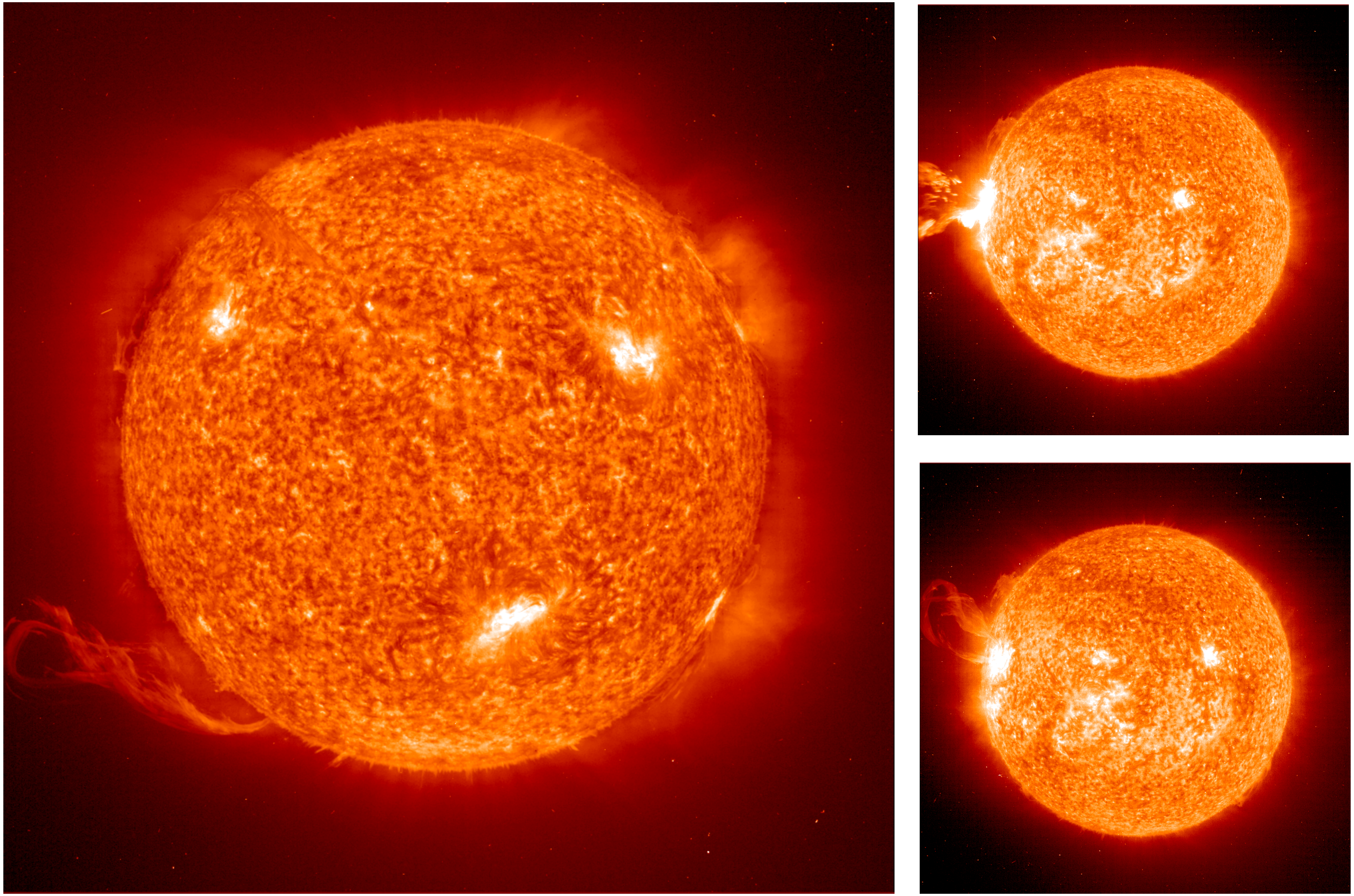
Übersicht

WAS?	- Was sind Sterne?
WO und WANN?	- Wo und wann bilden sich Sterne?
WIE?	- Wie beobachtet man Sternbildung? - Wie bilden sich Sterne?
→	Theorie der turbulenten Sternentstehung
BEISPIELE	- Sternbildung in Sonnennähe - Sternbildung im frühen Universum

Phänomenologie

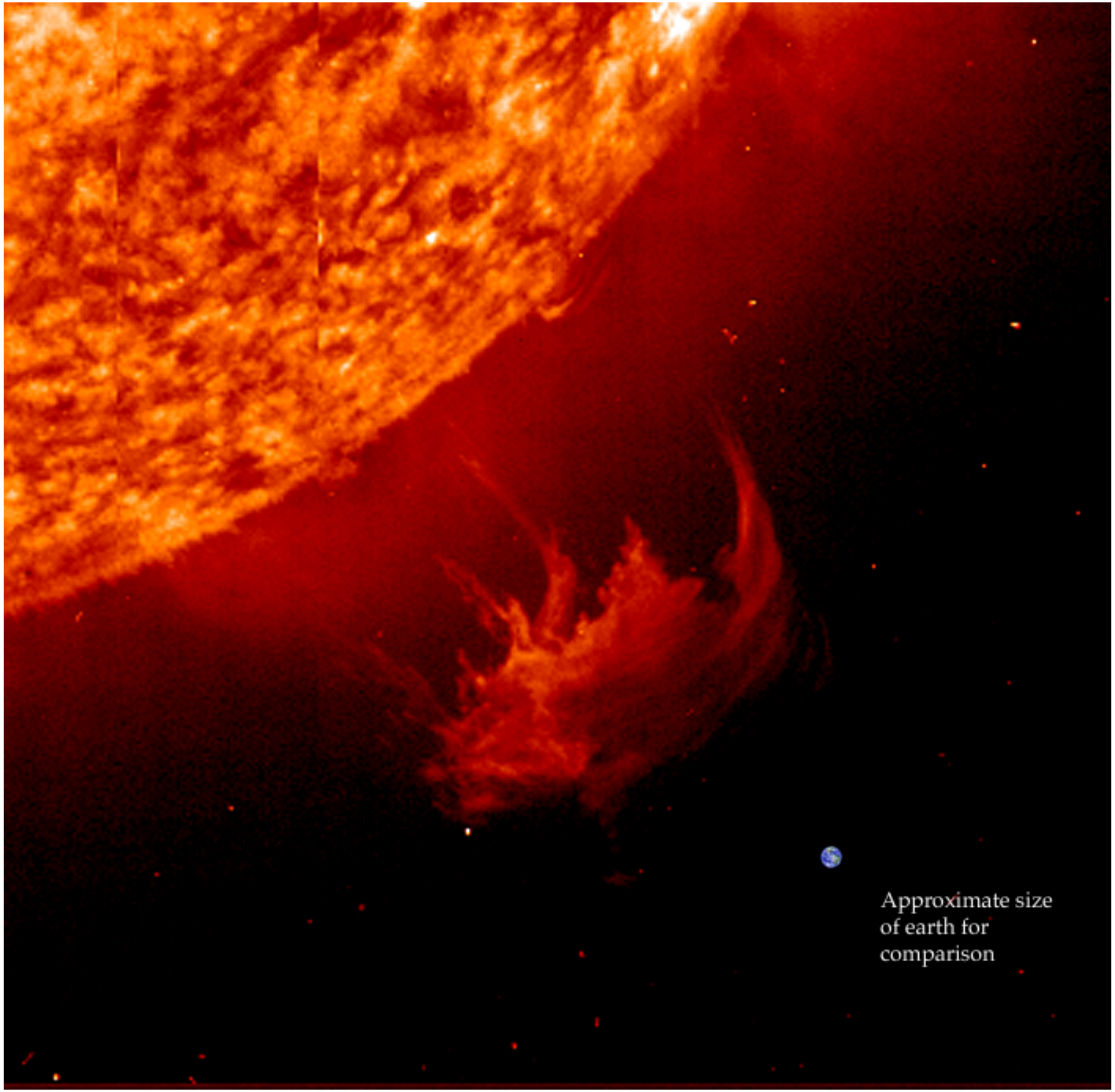
Theorie

WAS?

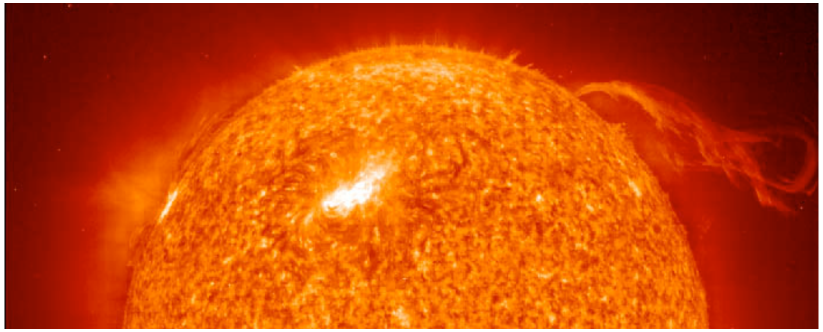


Unsere Sonne in verschiedenen Aktivitätsphasen

Größenvergleich: Sonne - Erde



Sterne: die Sonne



Eigenschaften der Sterne
(Stellare Zustandsgrößen):

Unsere Sonne ⊙ als Referenzstern

Radius	$R_\odot$	696 000 km	
Masse	$M_\odot$	$1{,}989 \times 10^{30}$ kg	
Leuchtkraft	$L_\odot$	$3{,}86 \times 10^{26}$ W	
effektive Temperatur	T_{eff}	5800 K	(Oberfläche)
Zentraltemperatur	$T_{zentral}$	15×10^6 K	
Alter	$t_\odot$	4.5×10^9 a	

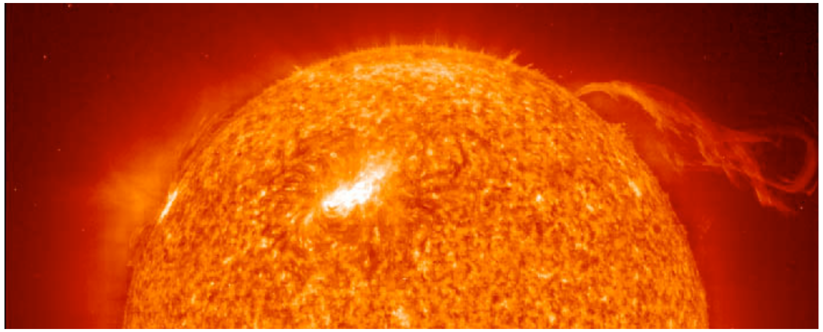
auf der Erde:
Solarkonstante
1.37 kW/m^2

Spektraltyp	G2
Leuchtkraftklasse	V

chemische Zusammensetzung (Massenanteil)

73% Wasserstoff X
25% Helium Y
2% Metalle Z

Sterne: die Sonne



Eigenschaften der Sterne
(Stellare Zustandsgrößen):

Unsere Sonne ⊙ als Referenzstern

Radius	$R_\odot$	7×10^{10} cm
Masse	$M_\odot$	2×10^{33} g
Leuchtkraft	$L_\odot$	4×10^{33} erg/s
effektive Temperatur	T_{eff}	5800 K
Zentraltemperatur	T_{zentral}	15×10^6 K
Alter	$t_\odot$	1.7×10^{17} s

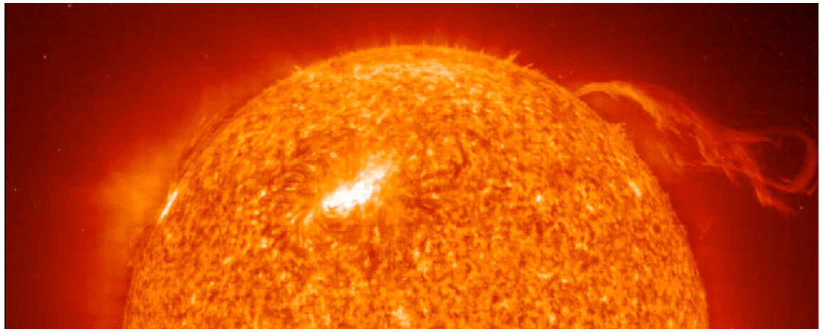
in cgs Einheiten

Spektraltyp	G2
Leuchtkraftklasse	V

chemische Zusammensetzung (Massenanteil)

73% Wasserstoff X
25% Helium Y
2% Metalle Z

Sterne: Statistische Charakeristika



Sterndichte in der Sonnenumgebung $n_* \approx 0{,}05\ \mathrm{pc}^{-3}$

Sterndichte im Orion-Trapezhaufen $n_* \approx 10^3\ \mathrm{pc}^{-3}$

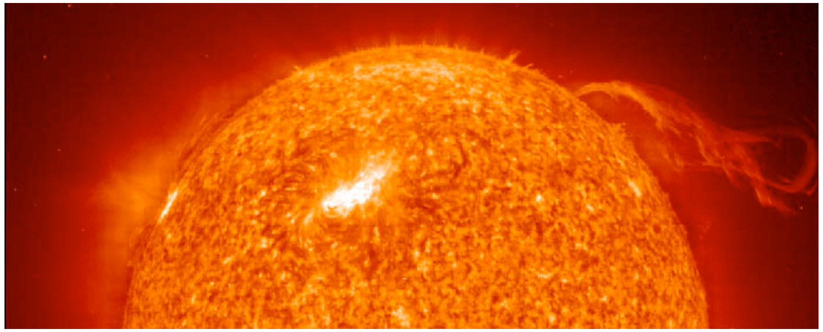
Sterndichte in Starburst-Regionen $n_* \approx 10^4\ \mathrm{pc}^{-3}$

$1\ \mathrm{pc} = 3{,}086 \times 10^{18}\ \mathrm{cm}$

Anzahl der Sterne in der Milchstraße: $\approx 10^{11}$

Sternentstehungsrate: $\approx 2\ M_\odot / \mathrm{Jahr}$

Sterne: Statistische Charakeristika



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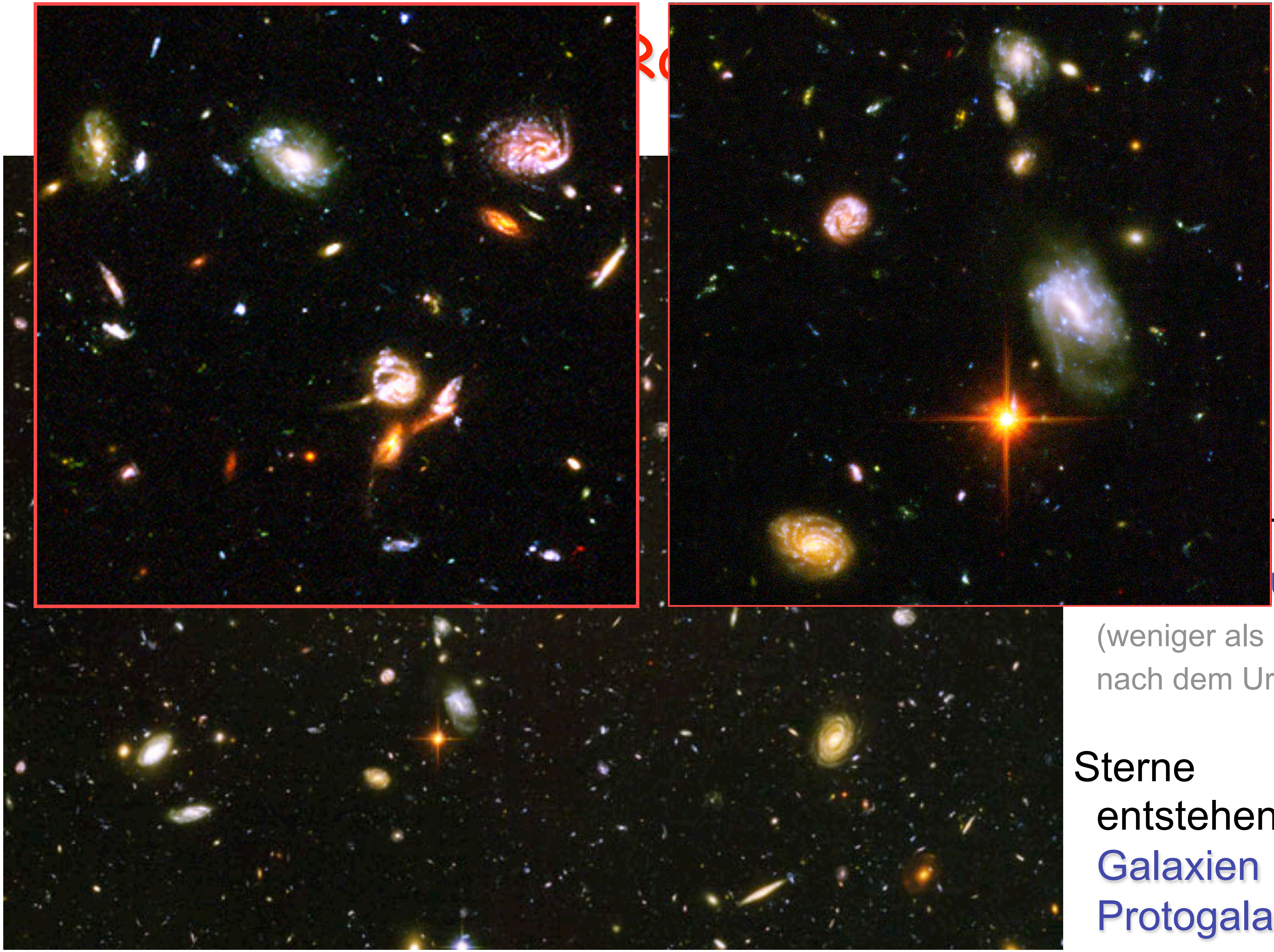
Anzahl der Sterne in der Milchstraße: $\approx 10^{11}$
Sternentstehungsrate: $\approx 2\ M_\odot / \mathrm{Jahr}$

Skalen:

Milchstraße ∅ $\approx$ 30 kpc $\approx 10^{23}$ cm $\approx$ 100.000 Lichtjahre
Sonnensystem (Pluto-Bahn) ∅ $\approx$ 80 AU $\approx 10^{15}$ cm $\approx$ 11 Lichtstunden
Erdbahn ∅ $\approx$ 2 AU $\approx 3 \times 10^{15}$ cm $\approx$ 17 Lichtminuten
Sonnendurchmesser ∅ $\approx 1{,}4 \times 10^{11}$ cm
Erddurchmesser ∅ $\approx 1{,}3 \times 10^{9}$ cm

WANN?

WO?



ng

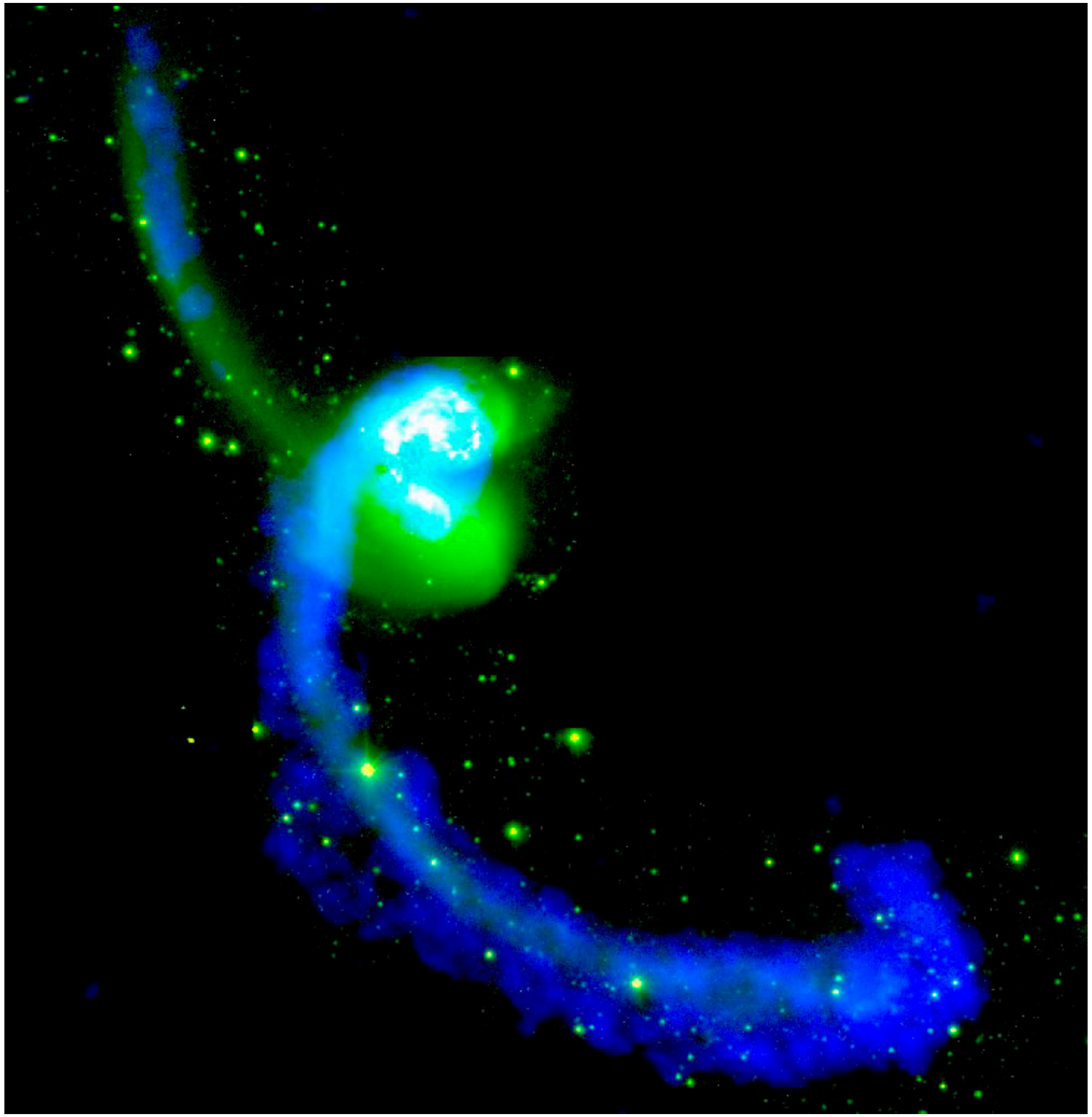
üh

(weniger als 1Ga nach dem Urknall)

Sterne entstehen in Galaxien und Protogalaxien

(Ausschnitt aus dem Hubble Ultra-Deep Field, von HST Webseite)

Sternentstehung in wechselwirkenden Galaxien

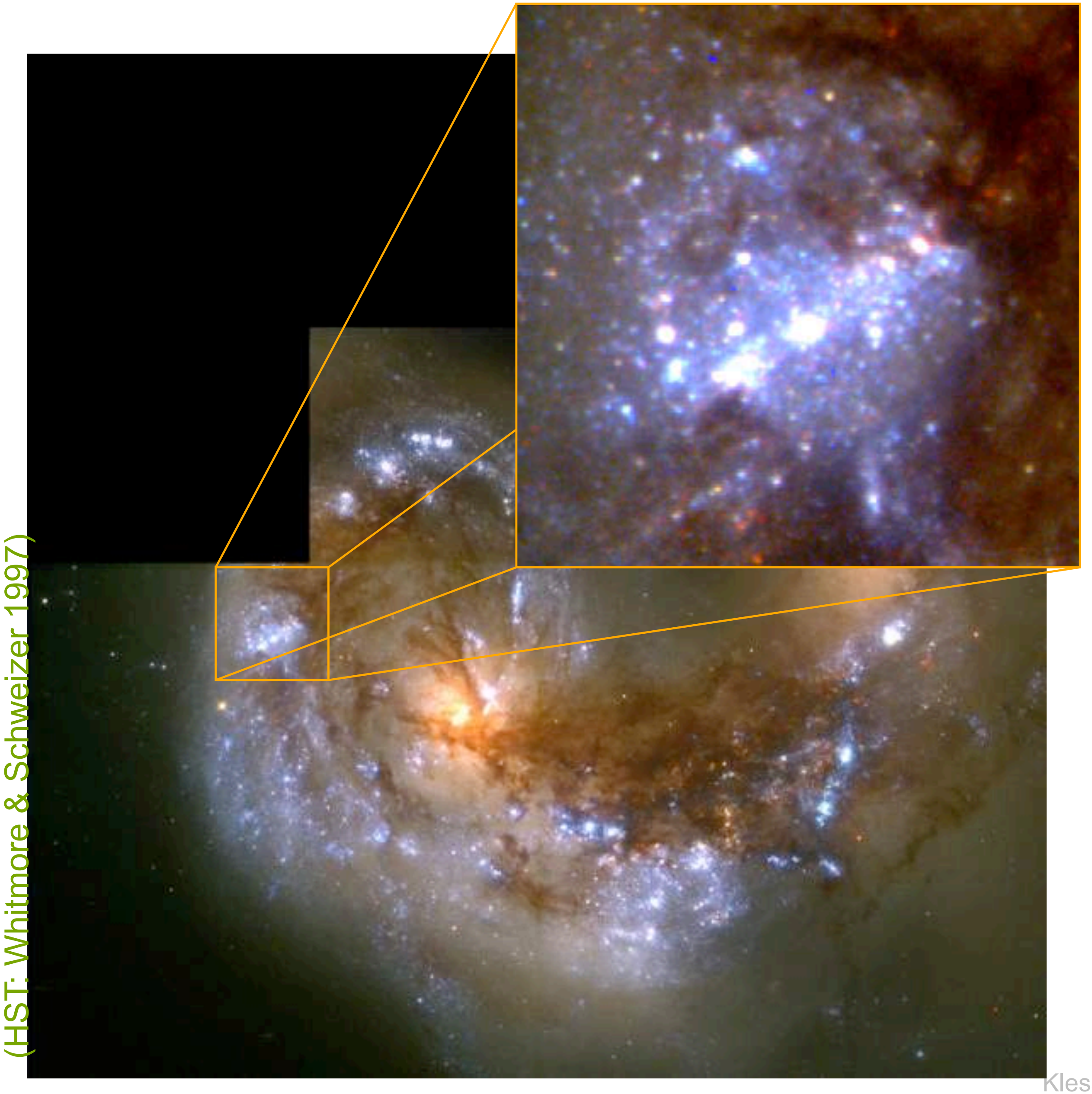


(Antennae Galaxie: Chandra Webpage)

Antennae Galaxie

- *NGC4038/39*
- *Entfernung: 19.2Mpc*
- *vis. Magn: 11.2*
- *optisch: weiß, grün*
- *radio: blau*

Sternentstehung in wechselwirkenden Galaxien



(HST: Whitmore & Schweizer 1997)

Antennae Galaxie

- Burst von Sternentstehung in wechselwirkenden Galaxien (dh. in stark gestörten Systeme)
- Sternbildung in Gezeitenarmen
- Großskalige, gravitative Bewegungen bestimmen SE
- Sterne entstehen in “Knoten” (dh. in Superhaufen)

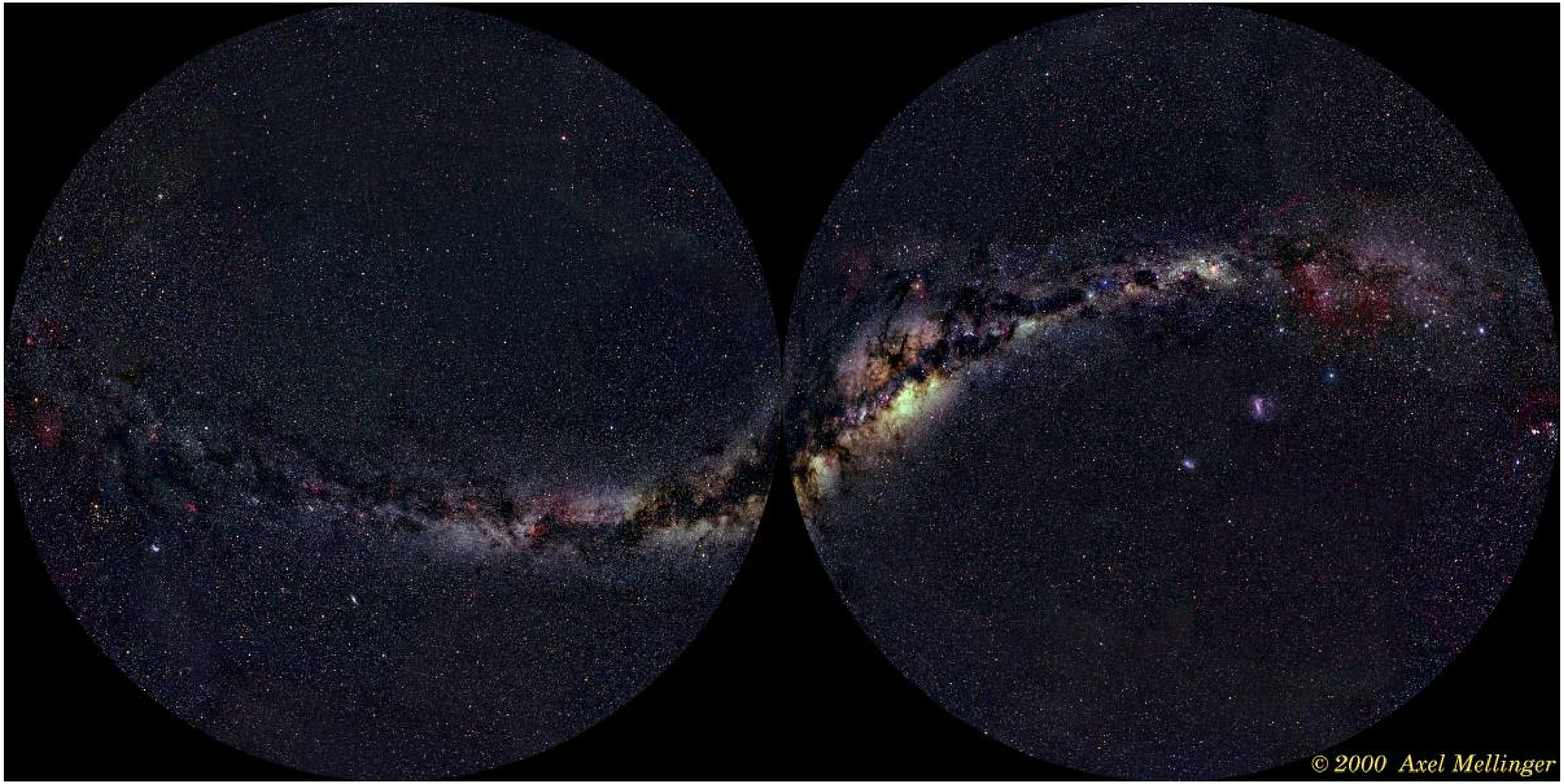
Sternentstehung in "normalen" Galaxien:

(from the Hubble Heritage Team)

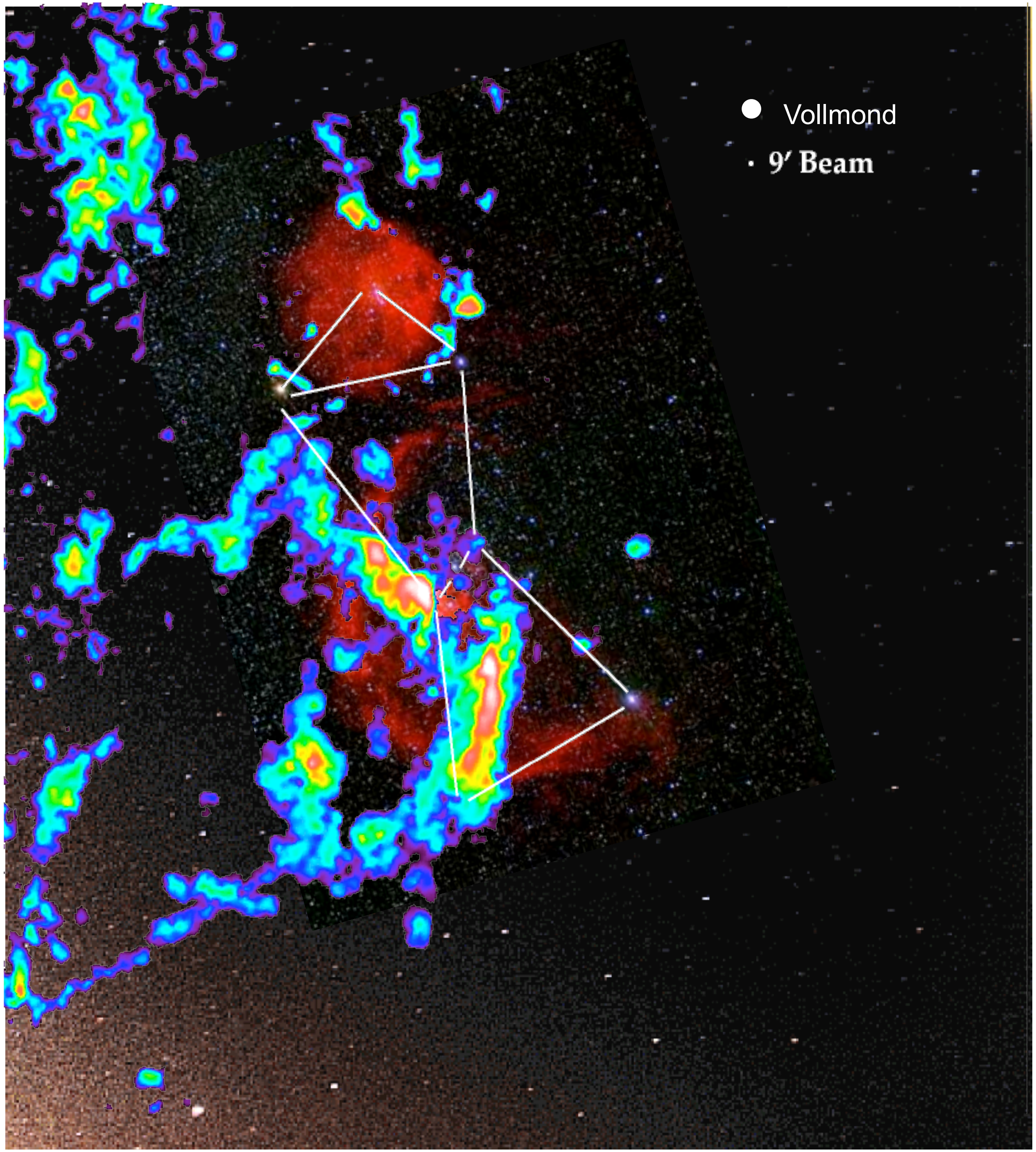
NGC4622

- Sternentstehung dauert bis *heute* an.
- Sternentstehung ist *immer* verbunden mit Wolken aus Gas und Staub.
- Staub- und Gasbänder findet man *häufig* (aber nicht immer) in Verbindung mit Spiralarmen
- Dasselbe gilt für die Sternentstehung.

Sternentstehung in der Milchstraße


© 2000 Axel Mellinger

Am Nachthimmel sieht man **Dunkelwolken** und **Sterne**:
Die hellsten Sterne sind *massereich* und daher *jung.*
→ Sternentstehung ist wichtig um beobachtete Struktur der Milchstraße zu verstehen.

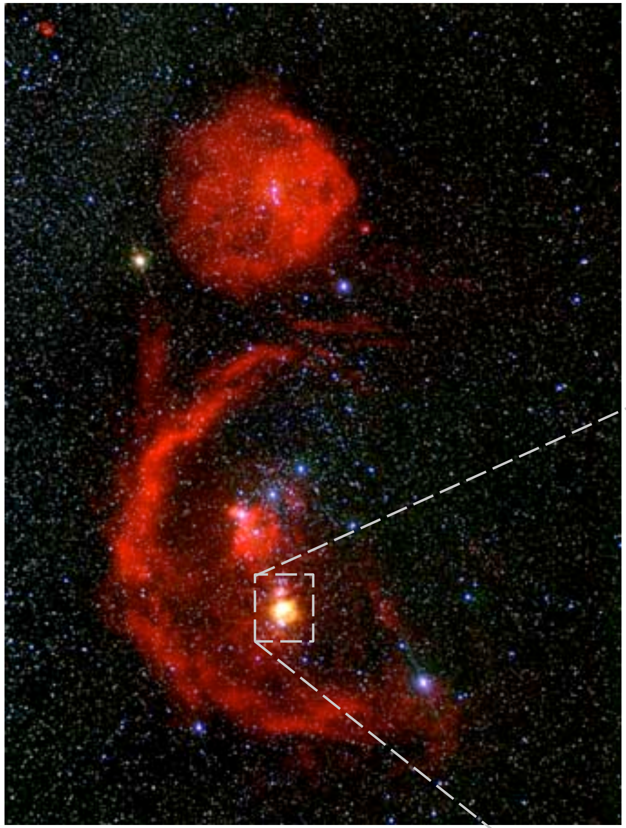


Stern-entstehung in Orion

Wir sehen

- *Sterne* (im sichtbaren Licht)
- Atomaren Wasserstoff (in Hα -- *rot*)
- Molekularen Wasserstoff H_2 (Radiostrahlung von Tracermolekül CO *farbcodiert*)

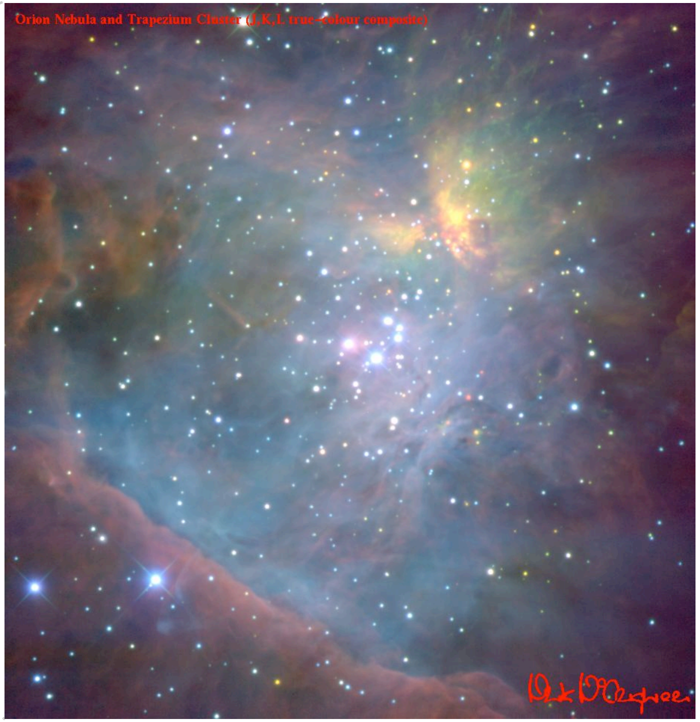
Lokales SE Gebiet: Trapezhaufen in Orion



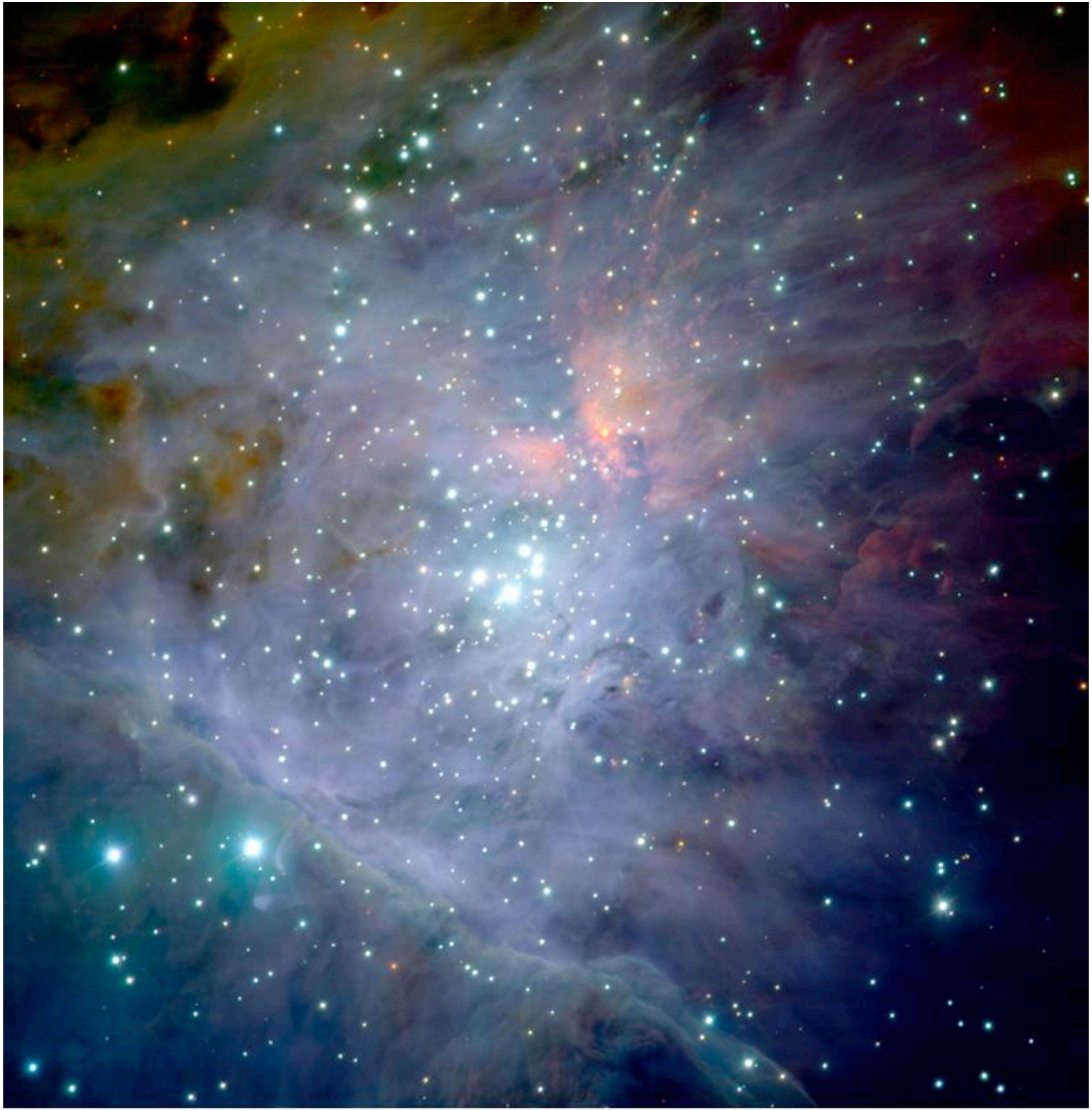
Sternbild Orion

Die Orionmolekülwolke ist die Geburtsstätte mehrer junger Sternhaufen.

Der Trapezhaufen ist noch ``eingebettet´´ und nur im IR Wellenbereich sichtbar. Der Haufen besteht aus ~2000 jungen Sternen.



Trapez-haufen



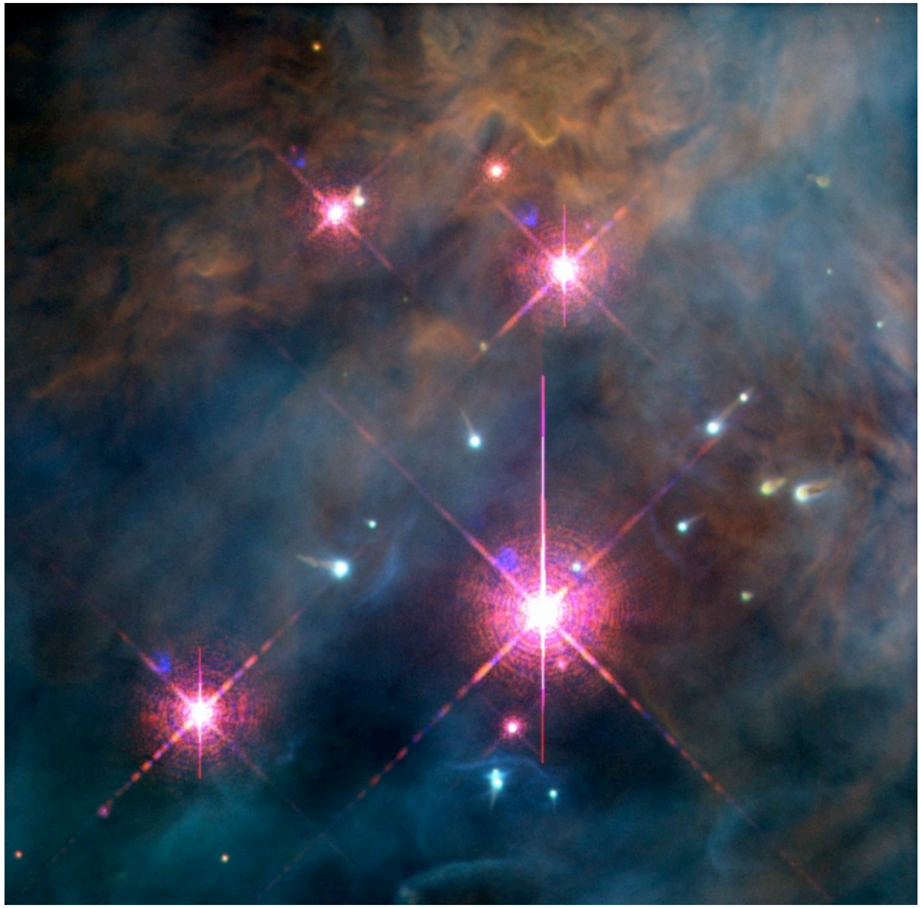
Trapezhaufen

(Detail)

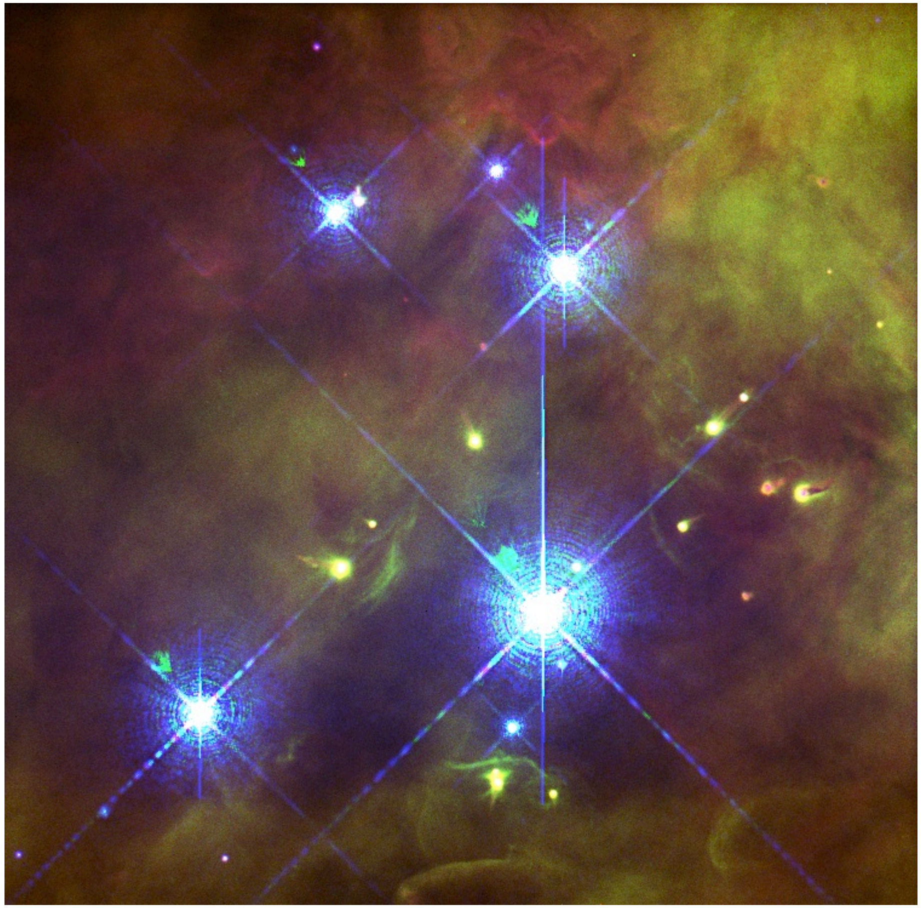
- Sterne entstehen in Haufen
- Sterne entstehen in Molekülwolken
- Rückkopplungseffekte sind wichtig

(Mehrfarbenaufnahme in J,H,K: McCaughrean, VLT, Paranal, Chile)

Trapezhaufen: Zentralgebiet



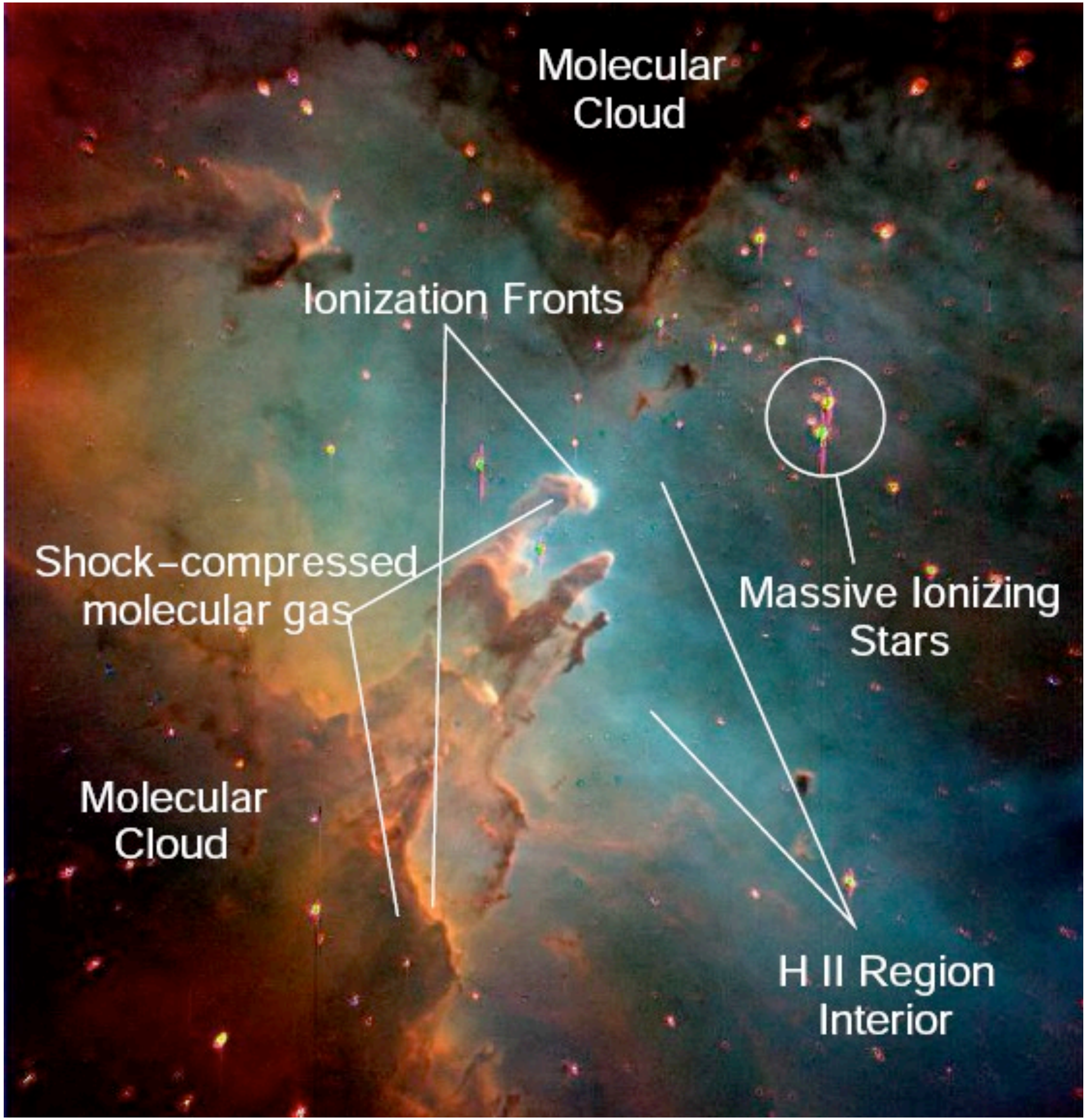
Der Großteil der ionisierenden Strahlung kommt vom Zentralstern **Θ1C Orionis**

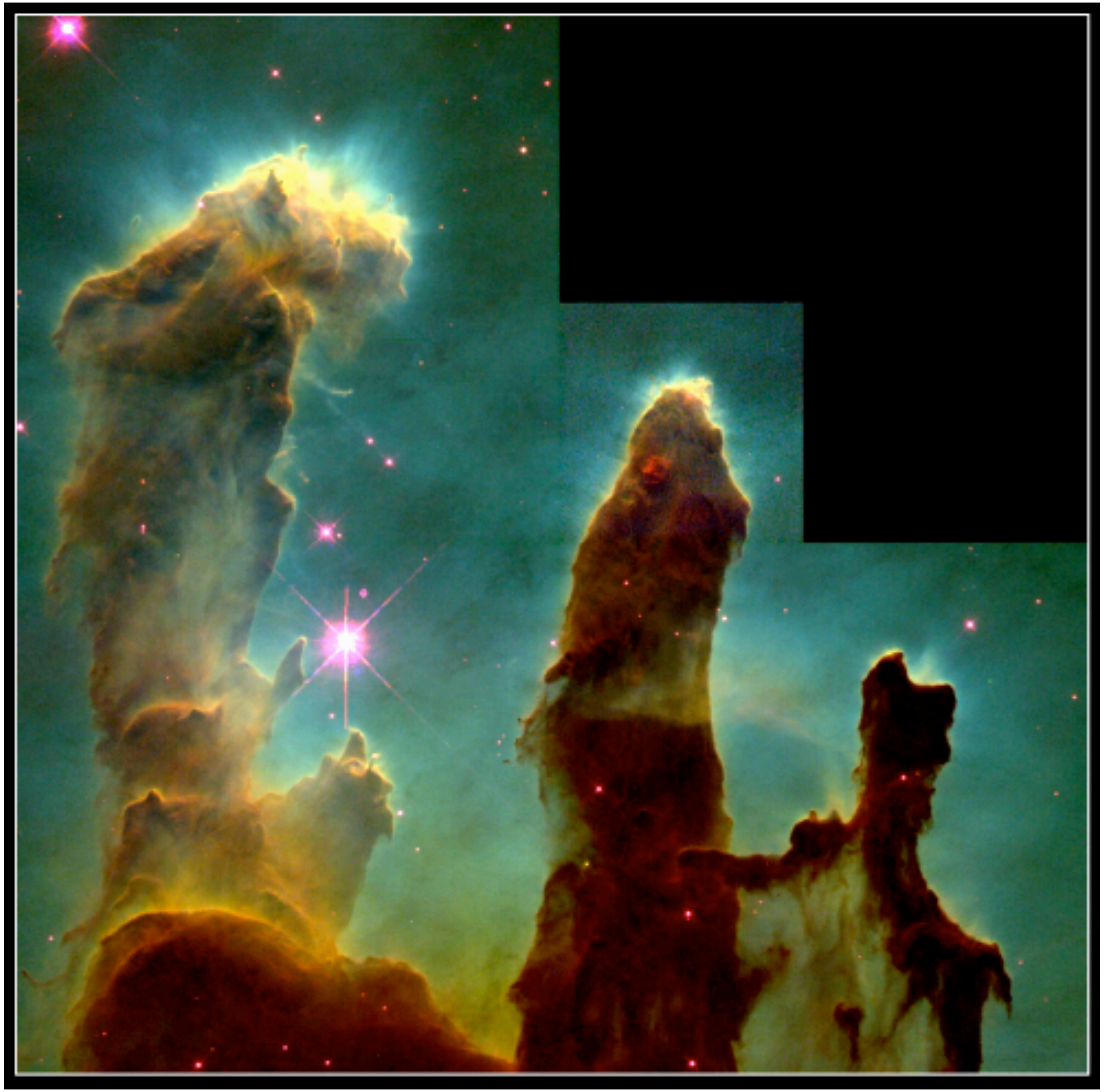


Proplyds: Verdampfende ``protoplanetare´´ Scheiben um junge massearme Protosterne

(Abbildung Doug Johnstone et al.)

alles in einem Bild





HST Aufnahme

Pillars of God (im Adlernebel): Entstehung kleiner Gruppen junger Sterne in den ``Spitzen´´ der Gas- und Staubsäulen….

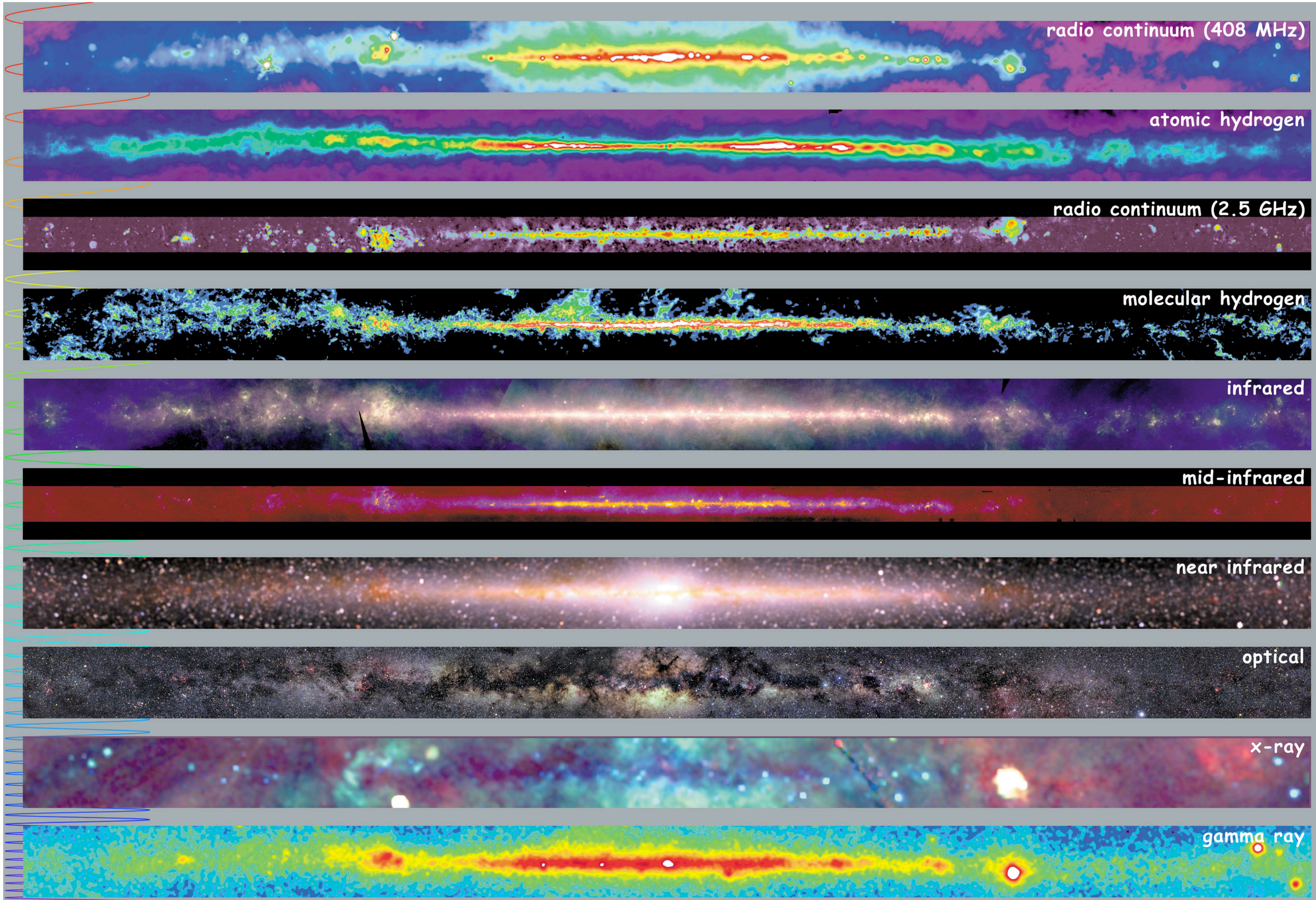
Aufnahme im
Infraroten.

(Andersen et al 2005)

IR Aufnahme mit dem ESO-VLT

Pillars of God (im Adlernebel): Entstehung
kleiner Gruppen junger Sterne in den ``Spitzen´´
der Gas- und Staubsäulen….

WIE?



http://adc.gsfc.nasa.gov/mw

Multiwavelength Milky Way

Wie beobachtet man SE?

Unterschiedliche Wellenlängen liefern unterschiedliche Informationen.

→Astronomen beobachten das gesamte elektromagnetische Spektrum

- Radiostrahlung: interstellares Gas (Linienemission -> Geschwindigkeitsinformation)
- sub-mm Bereich: Staub (thermische Emission)
- infrarot & optisch: Sterne
- Röntgen: Sterne (Coronae), Supernovae Reste (sehr heißes Gas)
- γ-Strahlung: Supernovae Reste (radioaktiver Zerfall, z.B. ^{26}Al), kompakte Objekte, Verschmelzen von Neutronensternen (Gamma-Ray-Burst)

Interstellare Materie: ISM

Häufigkeit bezogen auf 1.000.000 Wasserstoff-Atome

Element		Ordnungszahl	kosmischeHäufigkeit
Wasserstoff	H	1	1.000.000
Deuterium	$_1H^2$	1	16
Helium	He	2	68.000
Kohlenstoff	C	6	420
Stickstoff	N	7	90
Sauerstoff	O	8	700
Neon	Ne	10	100
Natrium	Na	11	2
Magnesium	Mg	12	40
Aluminium	Al	13	3
Silicium	Si	14	38
Schwefel	S	16	20
Calcium	Ca	20	2
Eisen	Fe	26	34
Nickel	Ni	28	2

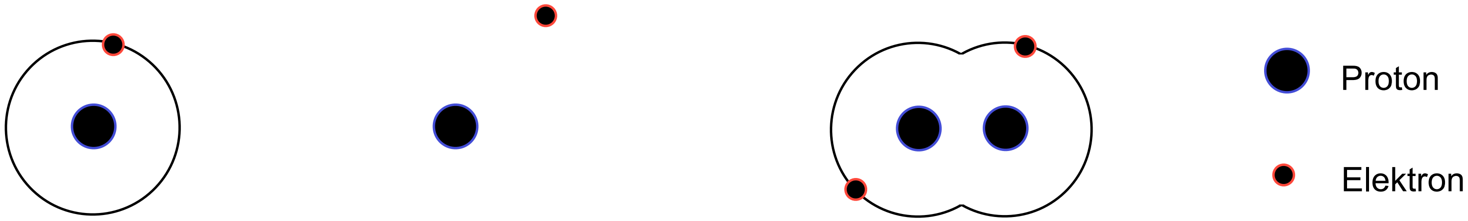
Wasserstoff ist das häufigste Element (mehr als 90% aller Atome). Im Vergleich zur kosmischen Häufigkeit sind manche Elemente im ISM seltener, d.h. abgereichert. Ein Teil ihrer Atome befinden sich nicht mehr in der Gasphase, sondern in Staubteilchen.

Phasen der ISM

Die Dominanz von Wasserstoff legt eine Klassifizierung der Regionen des ISM gemäß des Zustands von H nahe:

Ionisierter atomarer Wasserstoff HII (H^+)
Neutraler atomarer Wasserstoff HI (H)
Molekularer Wasserstoff H_2

Die jeweiligen Regionen bestehen nahezu zu 100% aus der entsprechenden Komponente, wobei, die Grenzgebiete zwischen HII, H und H_2 sehr dünn sind.



Phasen der ISM

Die Dominanz von Wasserstoff legt eine Klassifizierung der Regionen des ISM gemäß des Zustands von H nahe:

Ionisierter atomarer Wasserstoff	*HII* (H^+)
Neutraler atomarer Wasserstoff	*HI* (H)
Molekularer Wasserstoff	H_2

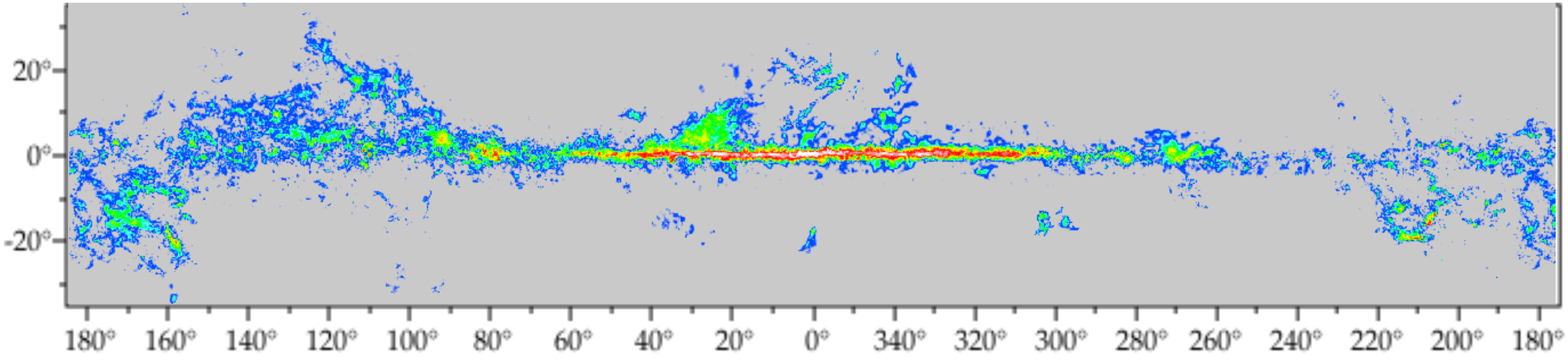
Ionisation

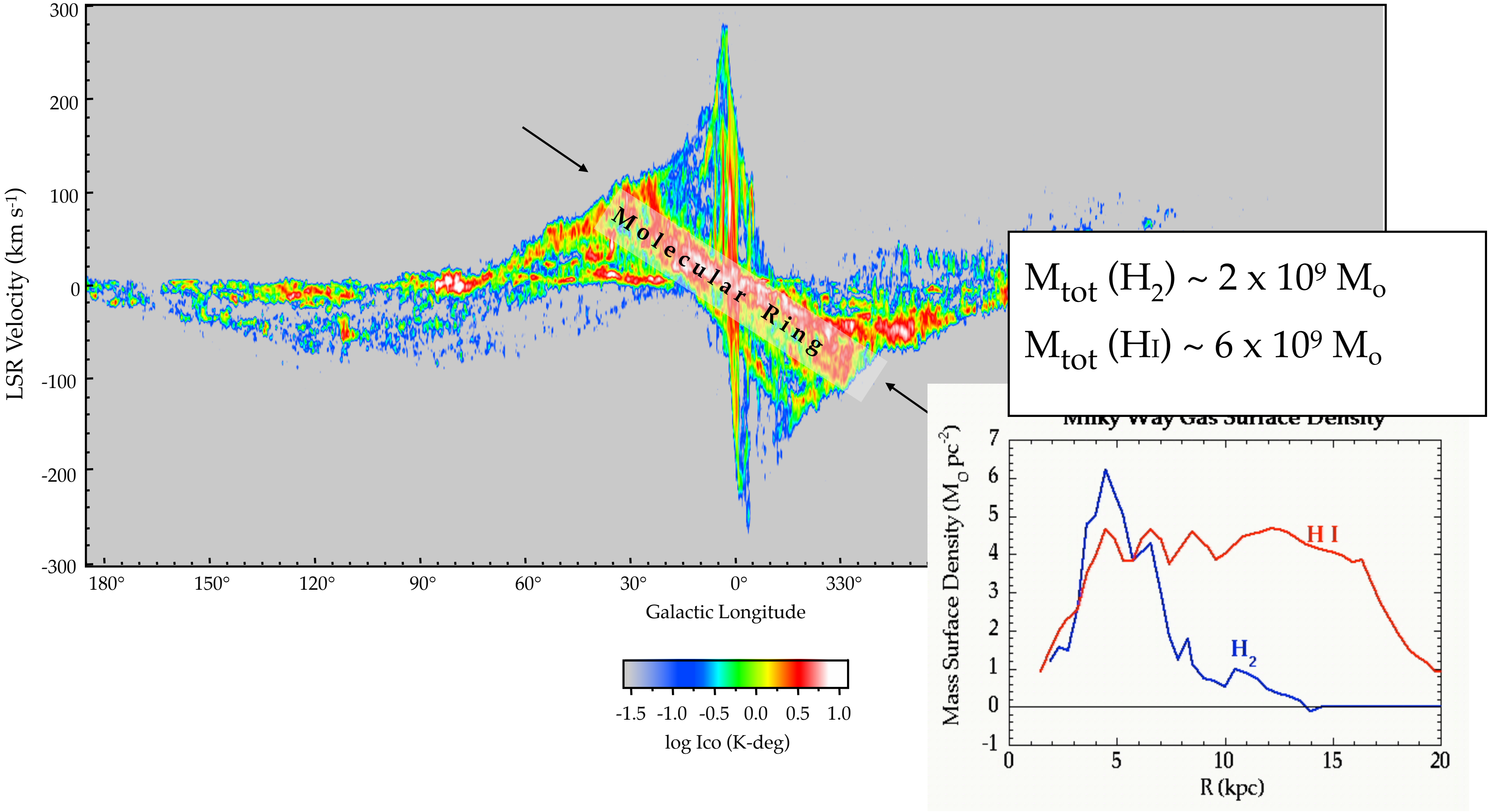
Phasenübergang

Die jeweiligen Regionen bestehen nahezu zu 100% aus der entsprechenden Komponente, wobei, die Grenzgebiete zwischen HII, H und H_2 sehr dünn sind.

Der größte Teil (ca. 80%) des ISM machen HI und HII-Gebiete niedriger Dichte aus. H_2 befindet sich in *Molekülwolken*, die oft von HII-Gebiete hoher Dichte begleitet werden. In der Milchstraße entstehen Sterne immer in Molekülwolken!

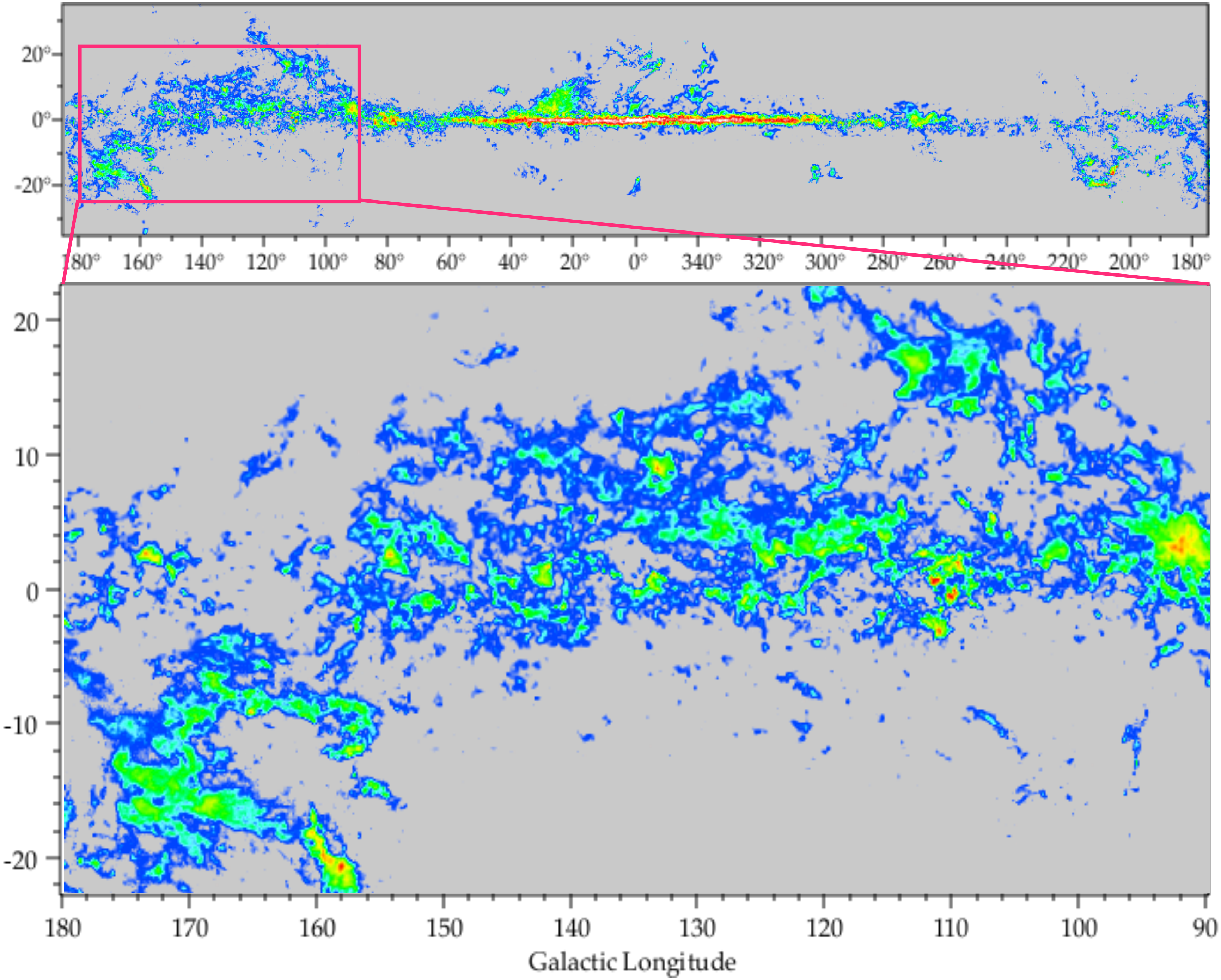
Data from Thomas Dame, CfA Harvard





M_{tot} (H_2) ~ 2 x 10^9 M_o

M_{tot} (HI) ~ 6 x 10^9 M_o



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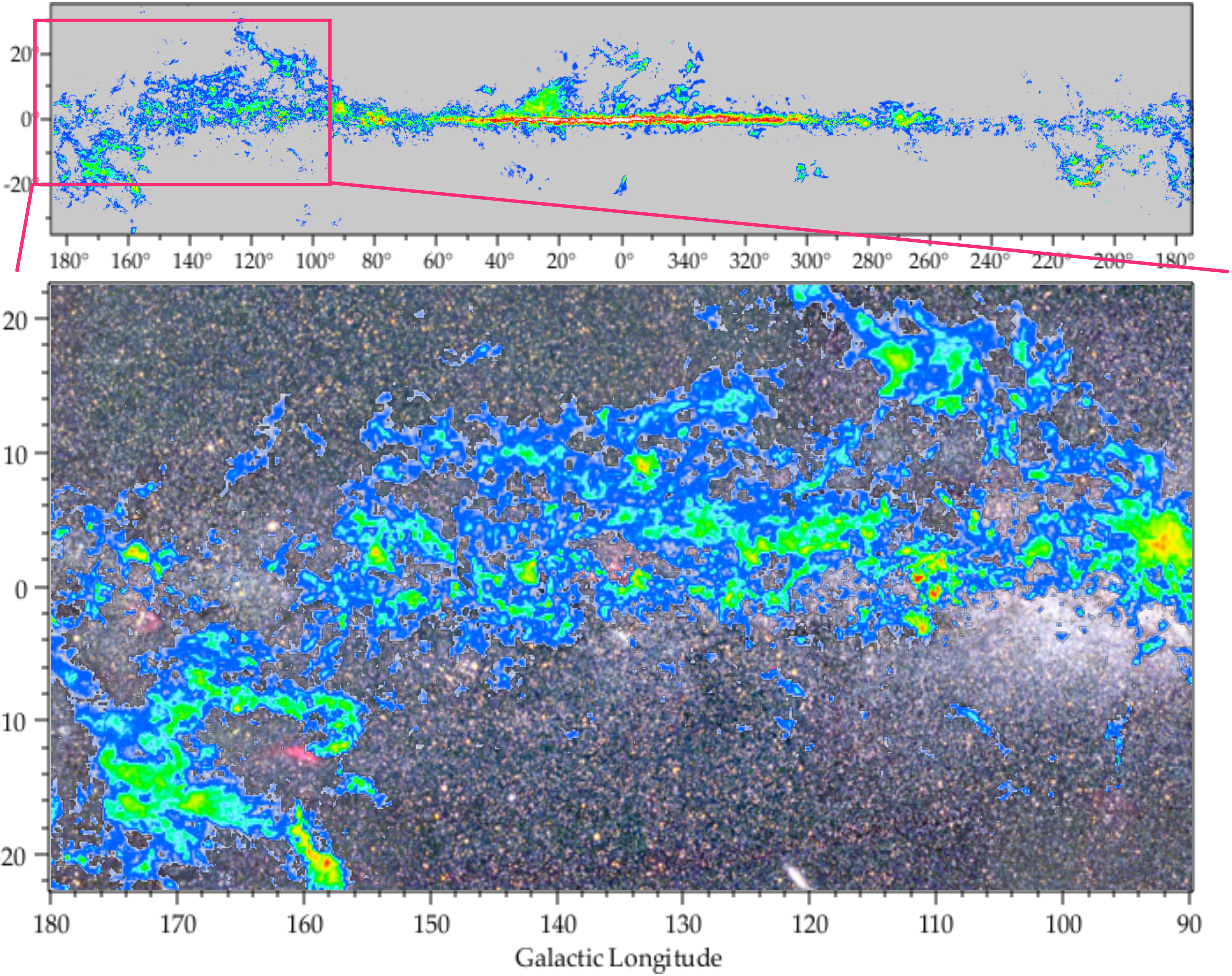

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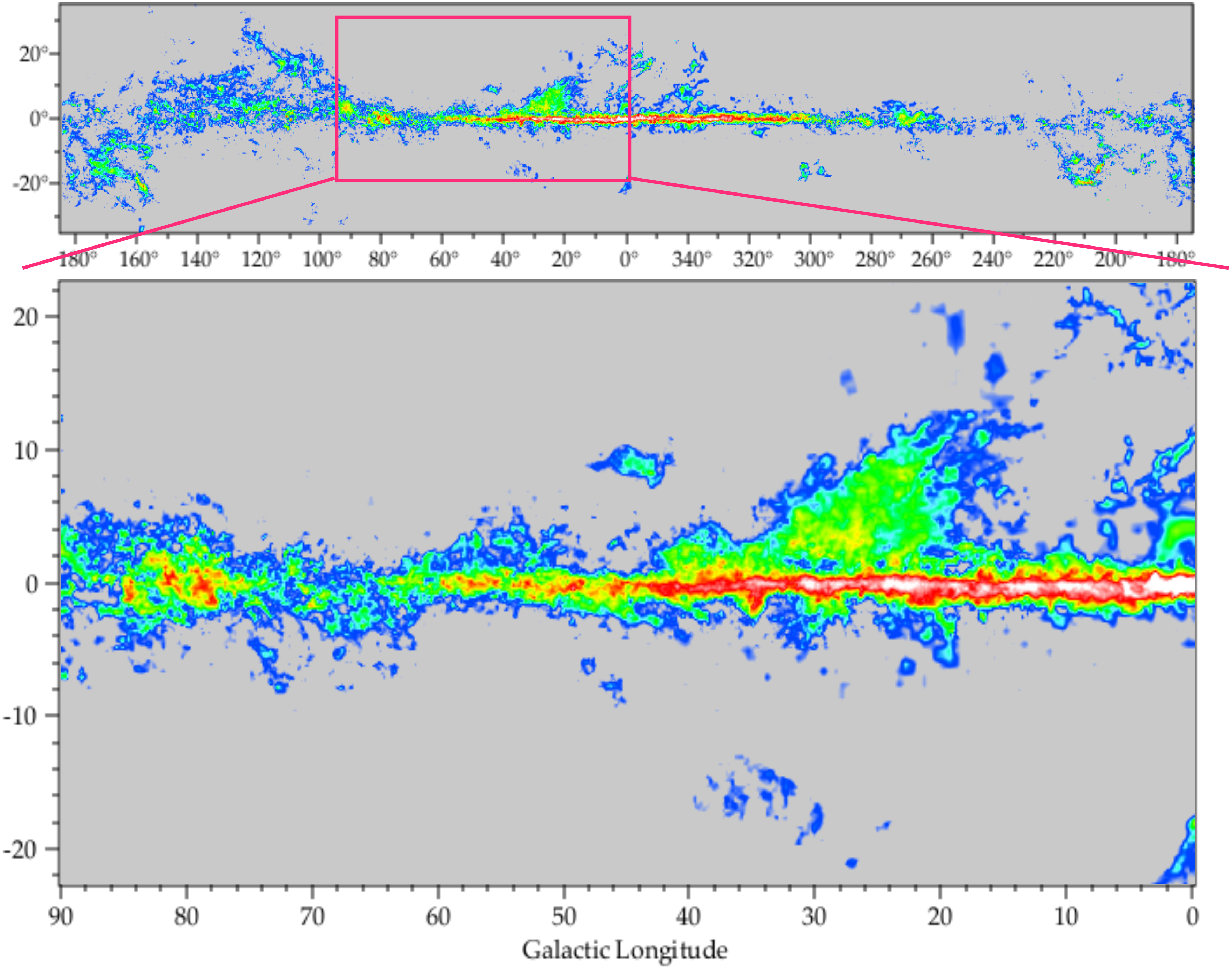
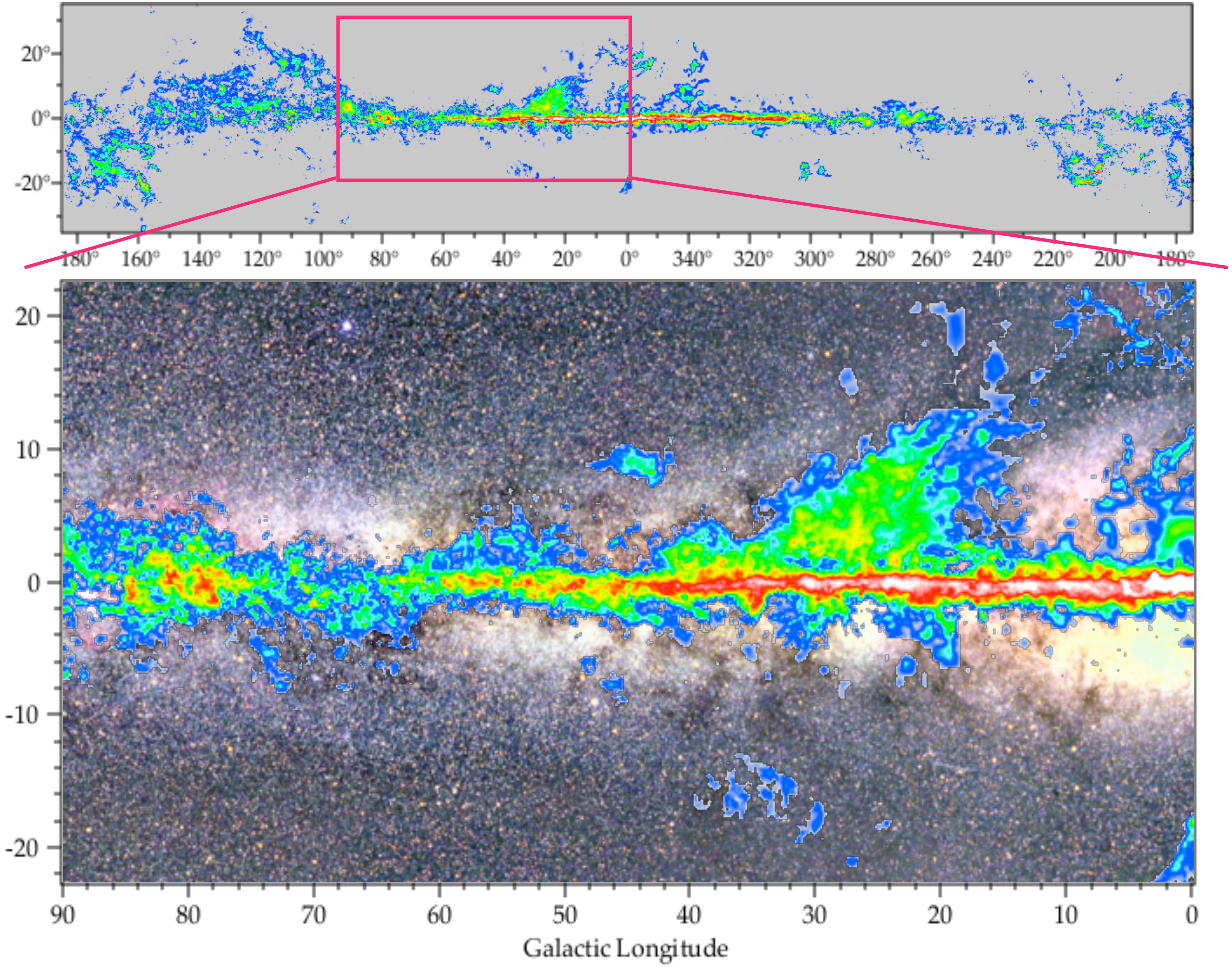


Data from Thomas Dame, CfA Harvard

Ralf Klessen: ISM lecture 25.09.2000



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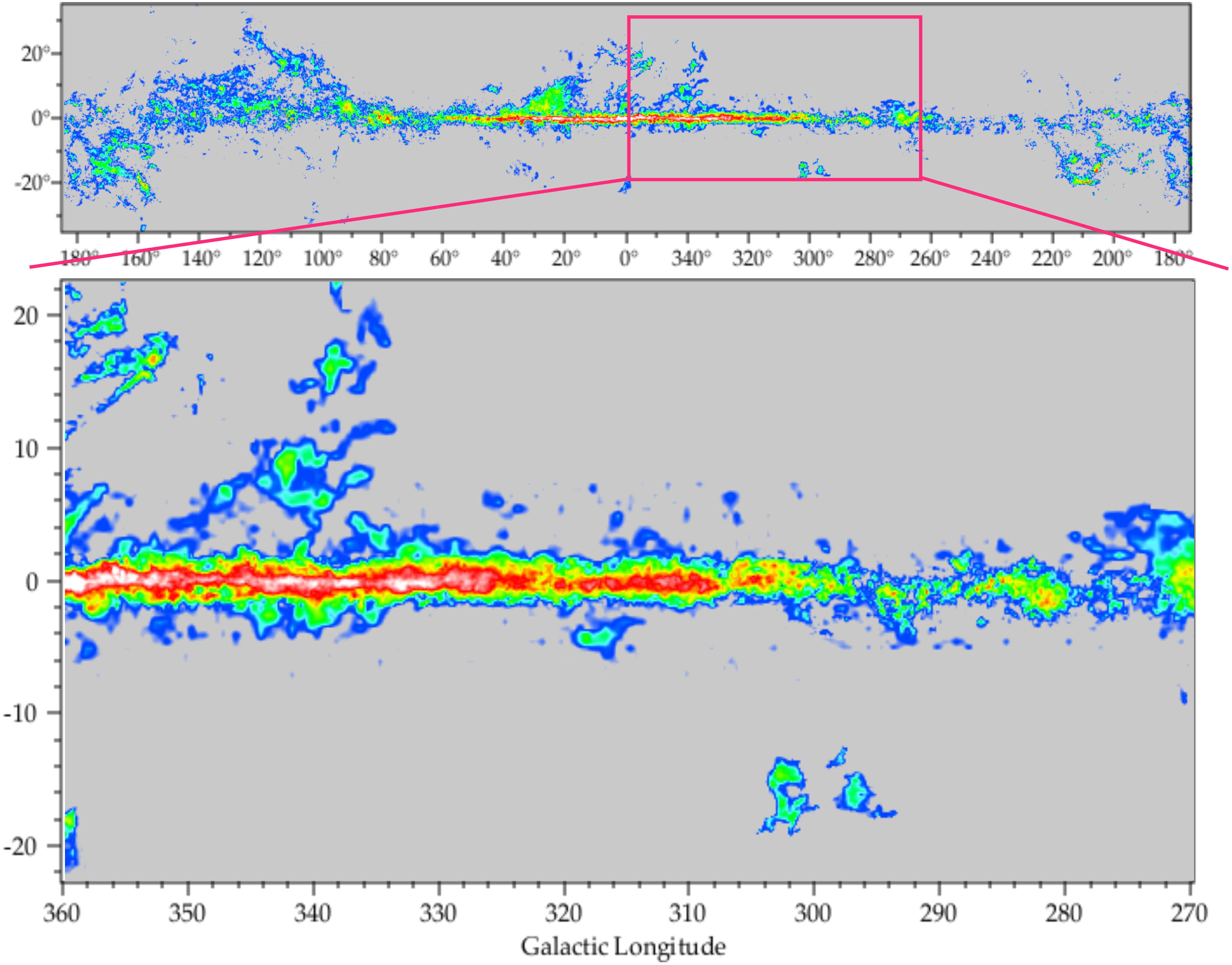

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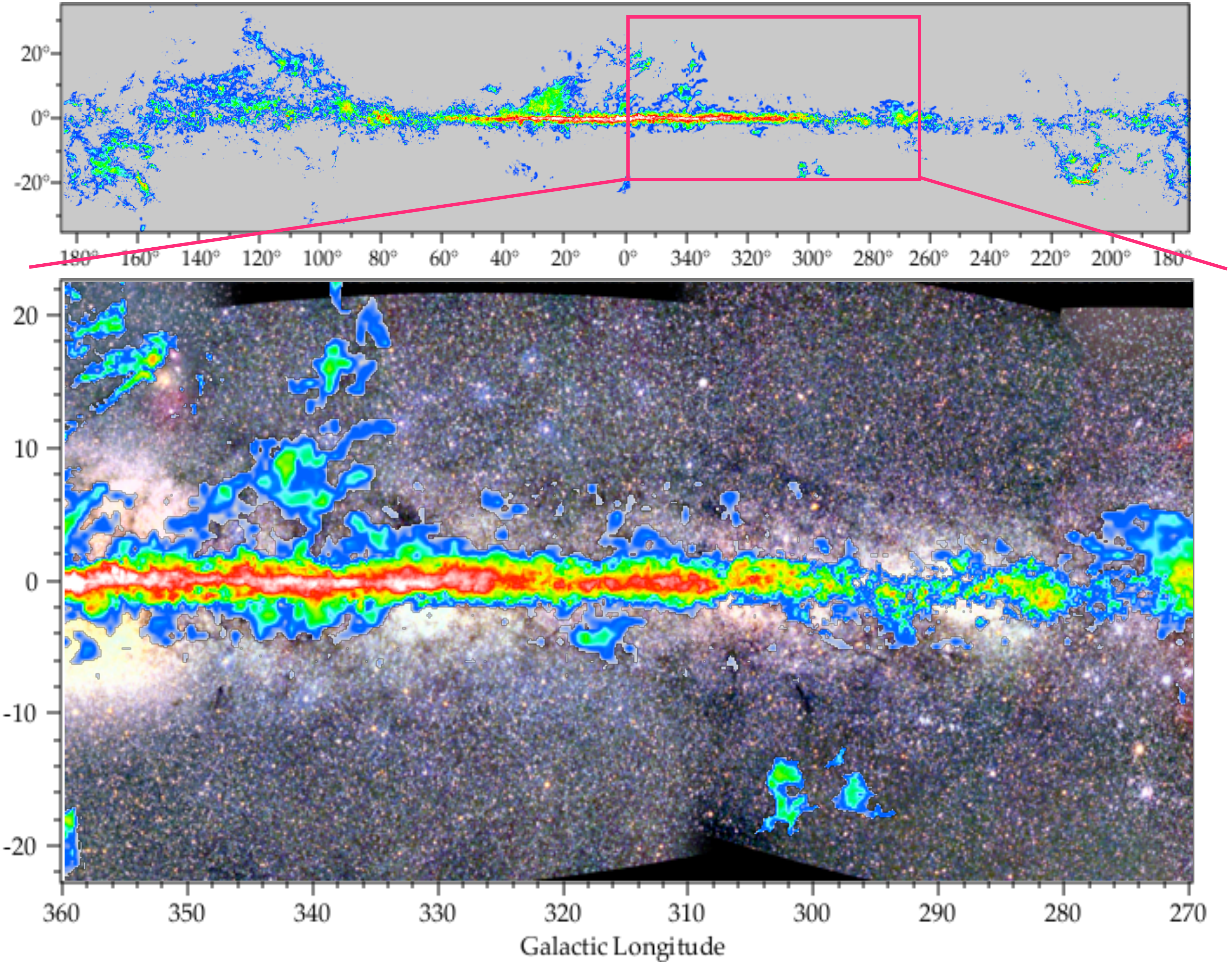

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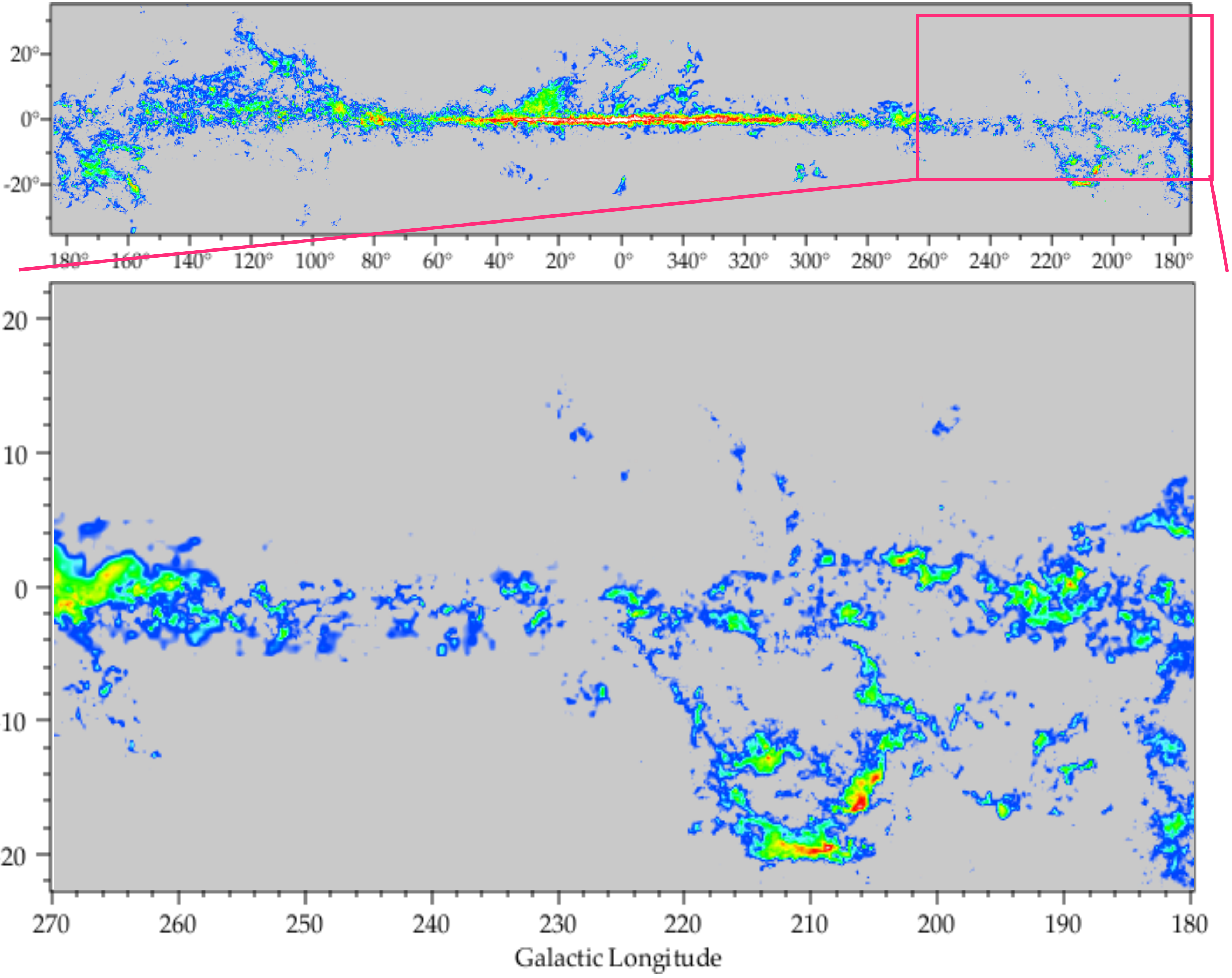

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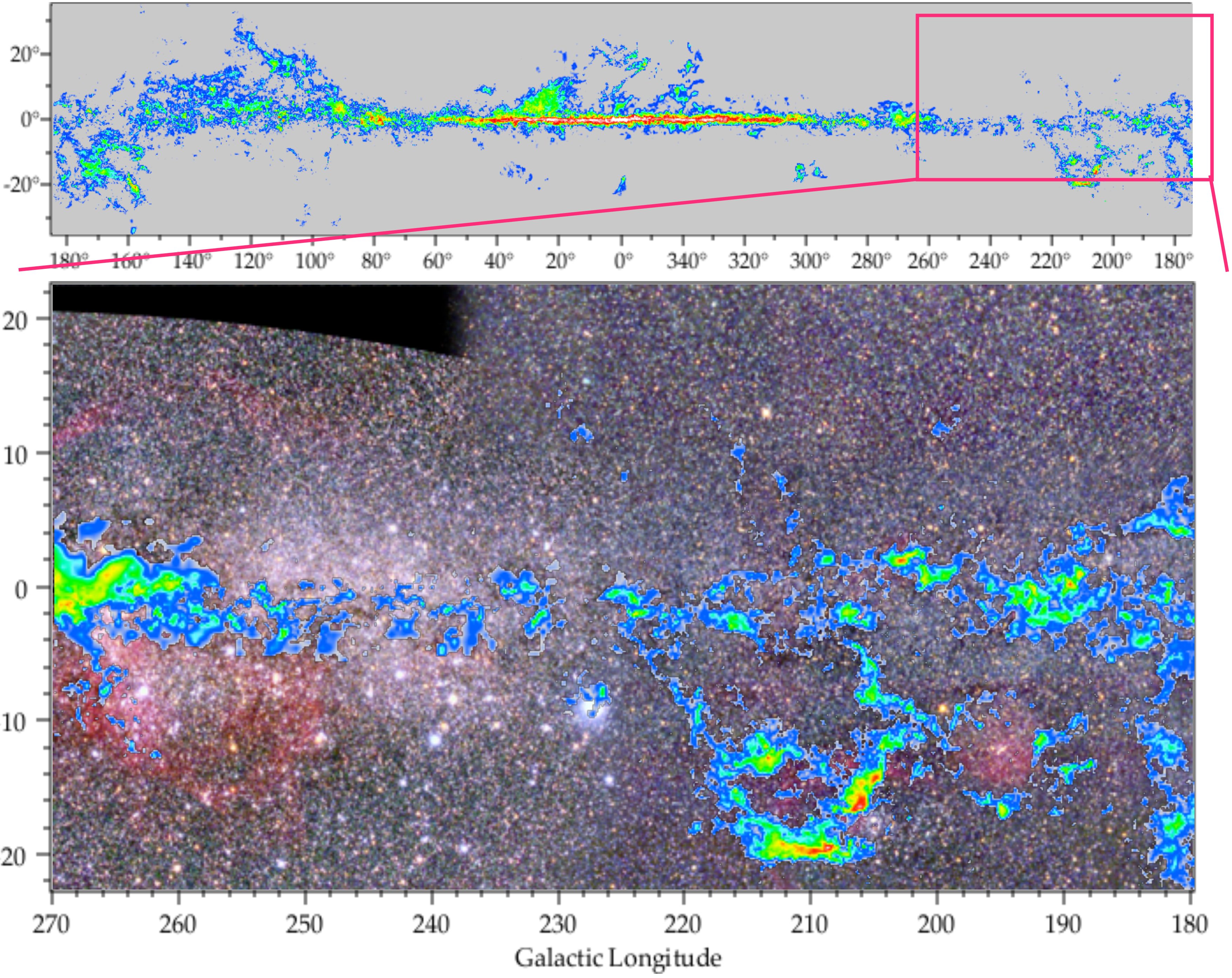

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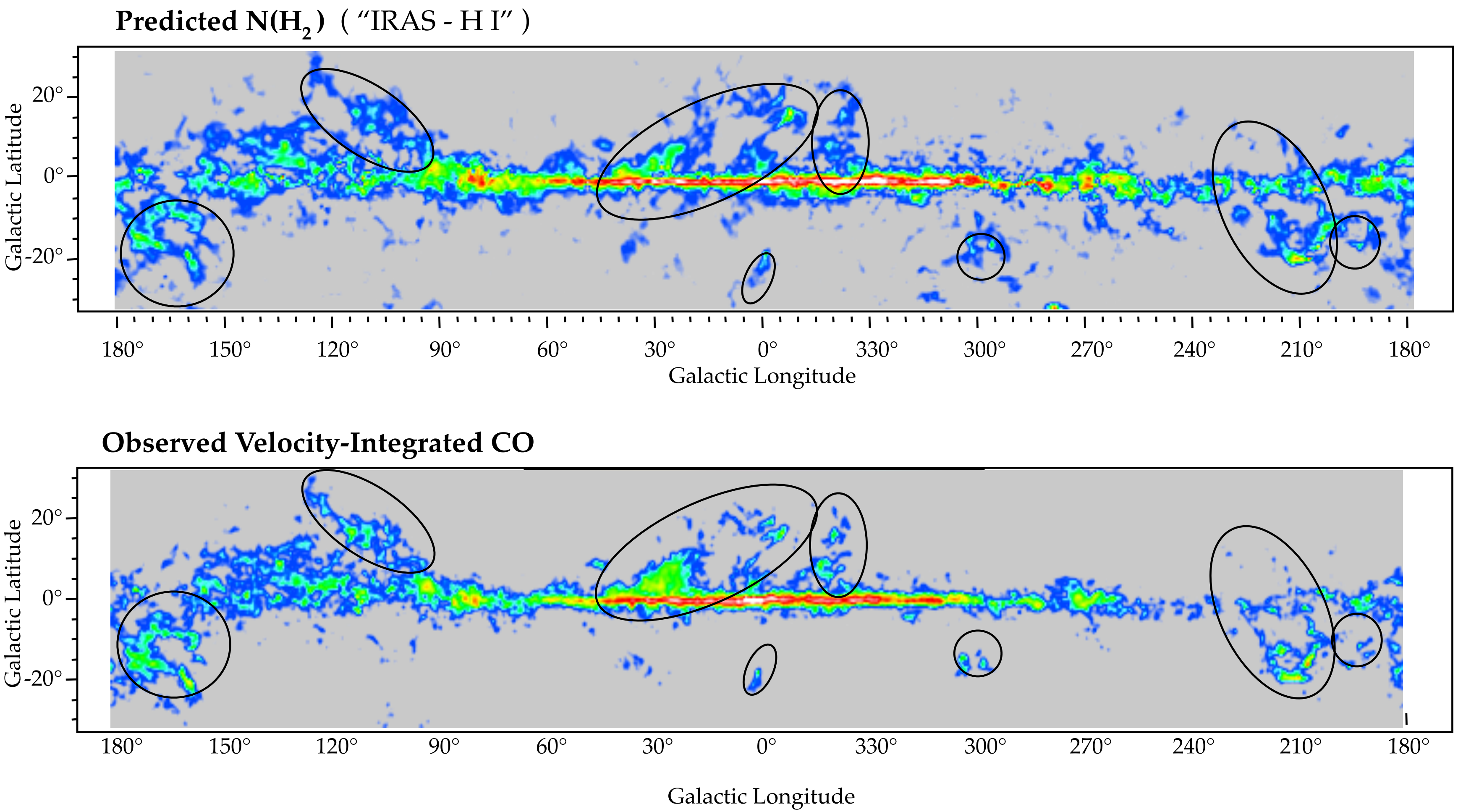


Polaris Flare
Cepheus

Ophiuchus | Lupus

Orion

Taurus

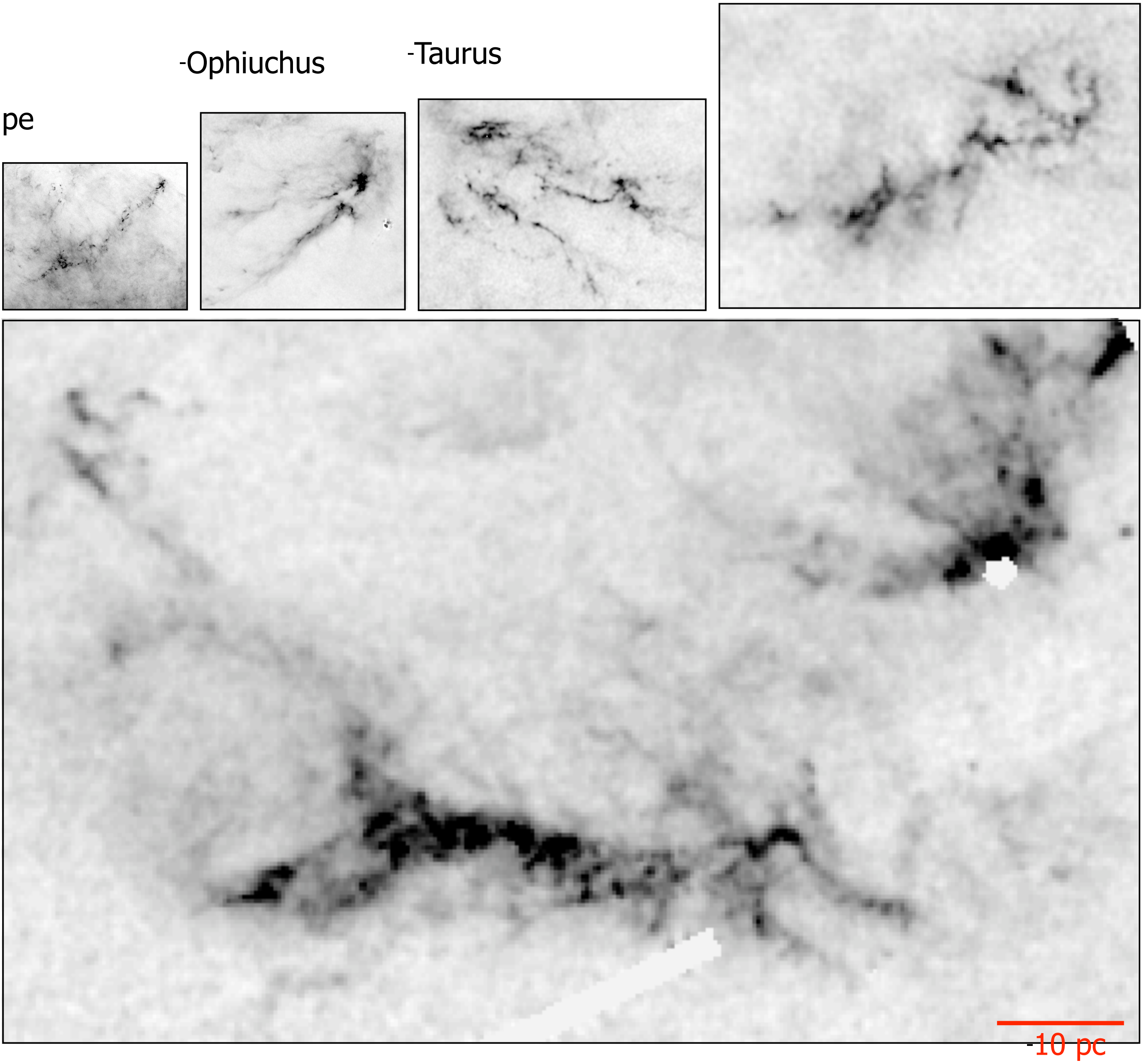
λ-Ori

RCrA

Chamaeleon

Data from Thomas Dame, CfA Harvard

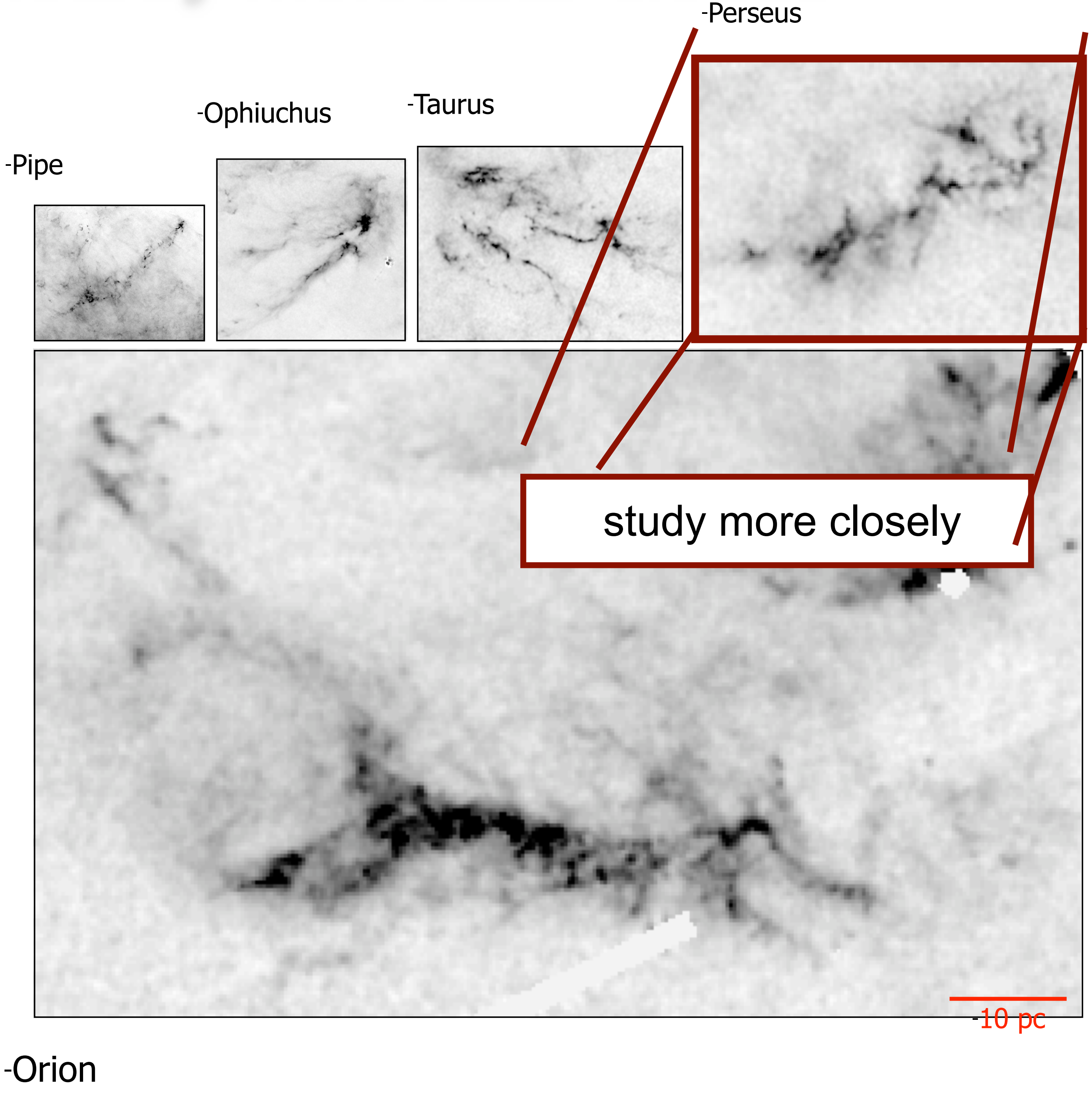
nearby molecular clouds



scales to same scale

(from A. Goodman)

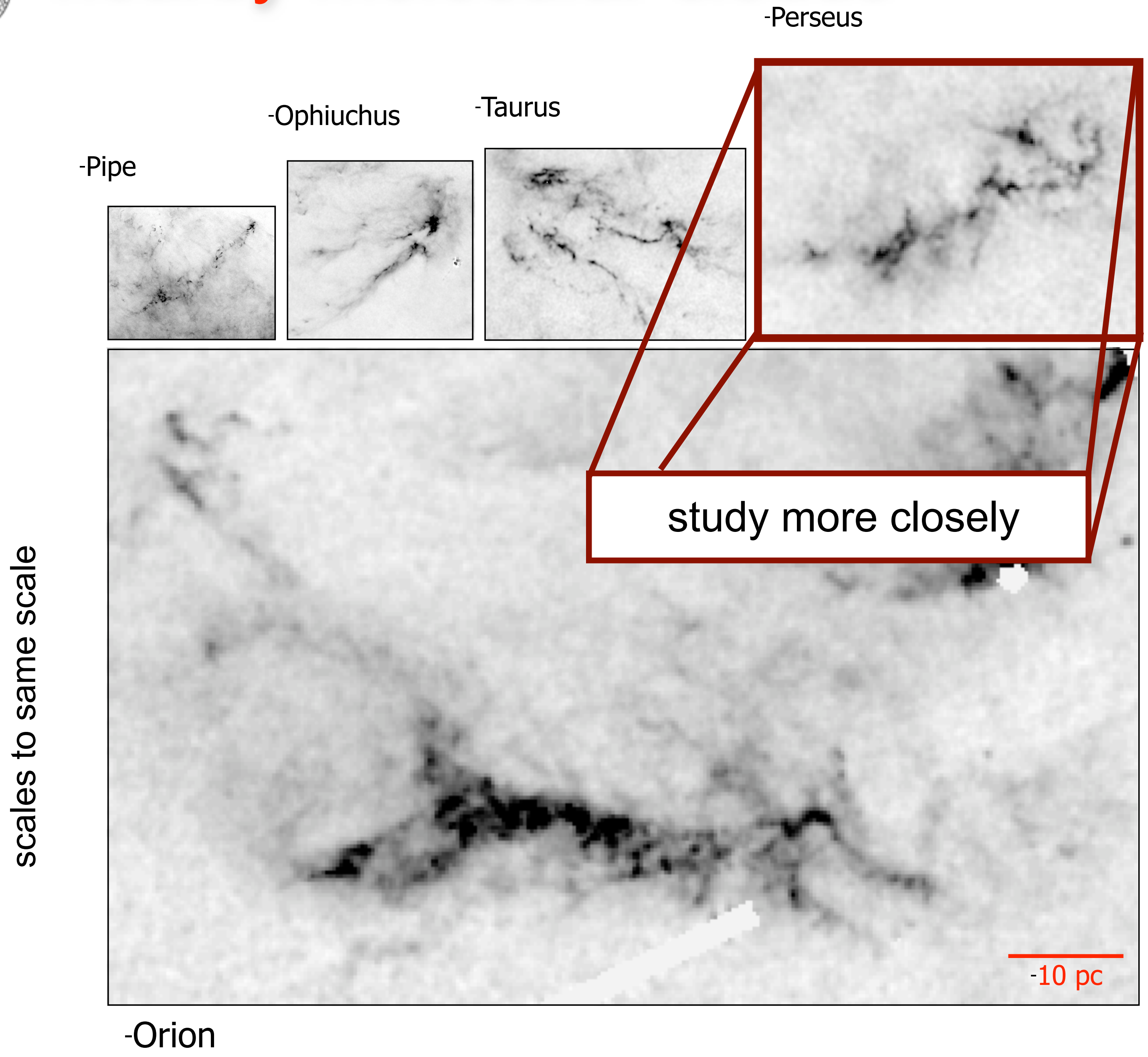
nearby molecular clouds



scales to same scale

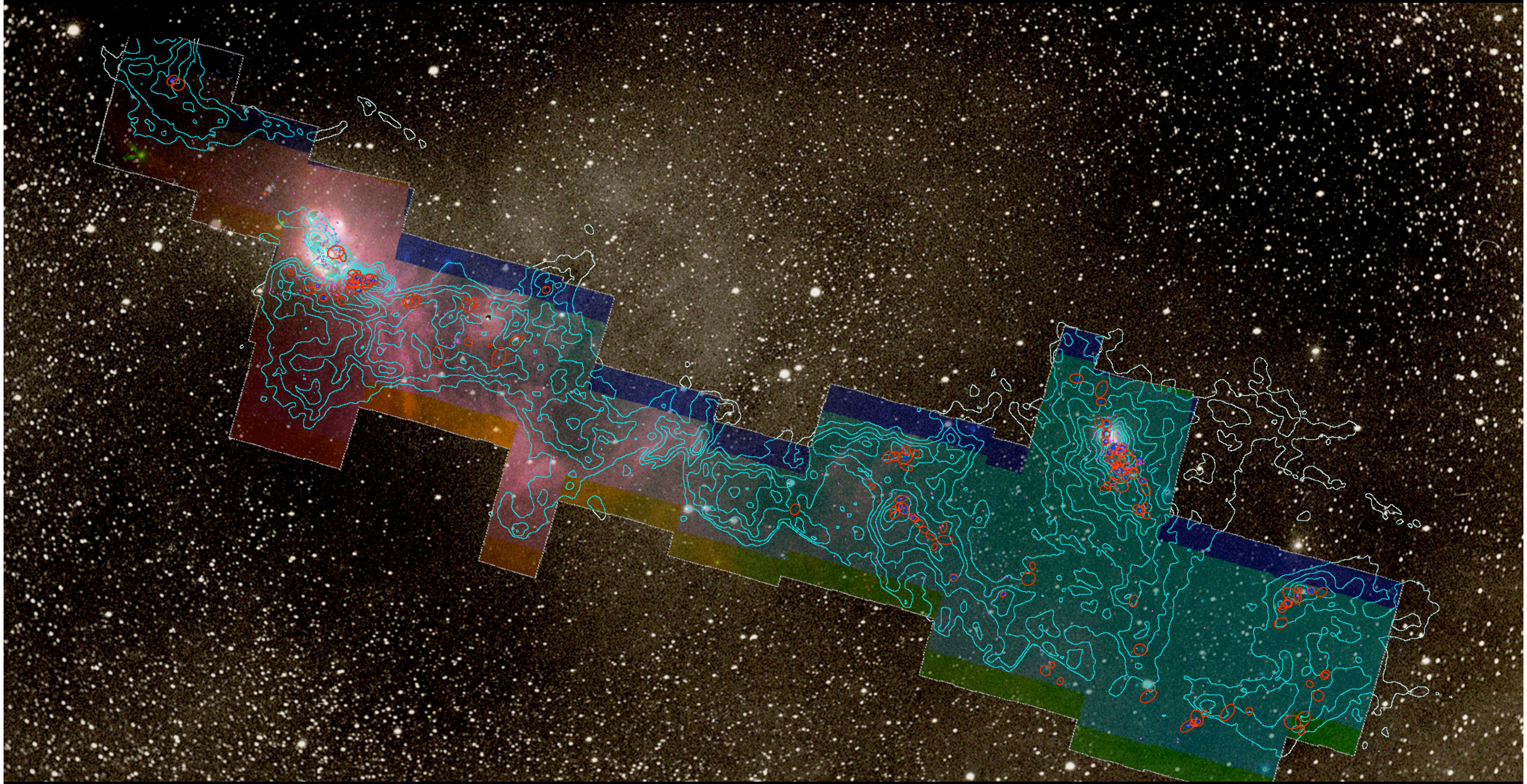
(from A. Goodman)

nearby molecular clouds



(from A. Goodman)

COMPLETE = COordinated Molecular Probe Line Extinction Thermal Emission Survey of Star-Forming Regions



COMPLETE Collaborators,
Summer 2008:

Alyssa A. Goodman (CfA/IIC)

João Alves (Calar Alto, Spain)

Héctor Arce (Yale)

Michelle Borkin (IIC)

Paola Caselli (Leeds, UK)

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Katherine Guenthner (CfA/Leipzig)

Mark Heyer (UMASS/FCRAO)

Doug Johnstone (HIA, Canada)

Jens Kauffmann (CfA/IIC)

Helen Kirk (HIA, Canada)

Di Li (JPL)

Jaime Pineda (CfA, PhD Student)

Erik Rosolowsky (UBC Okanagan)

Rahul Shetty (CfA)

Scott Schnee (Caltech)

Mario Tafalla (OAN, Spain)

COMPLETE

Perseus

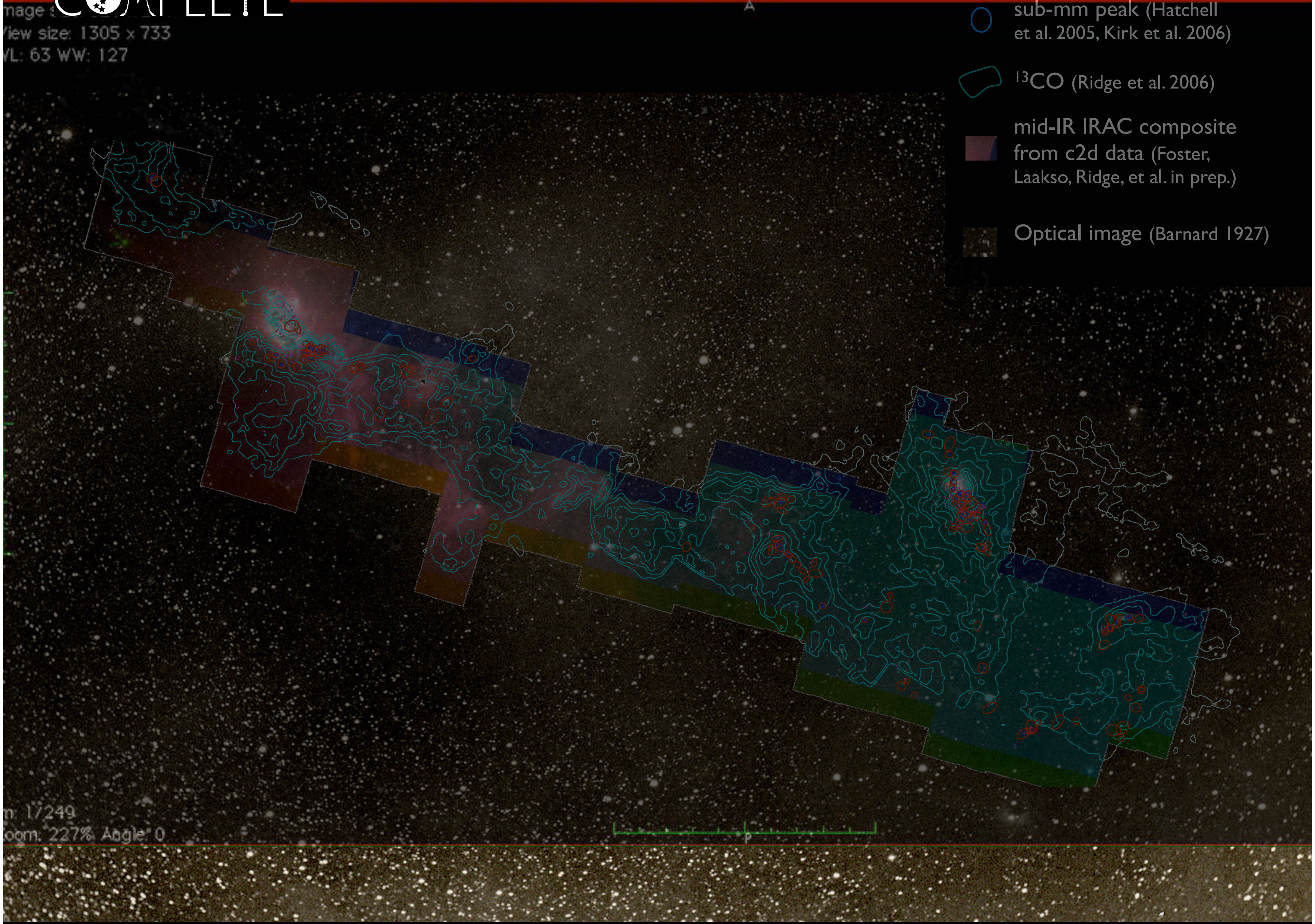
○ mm peak (Enoch et al. 2006)

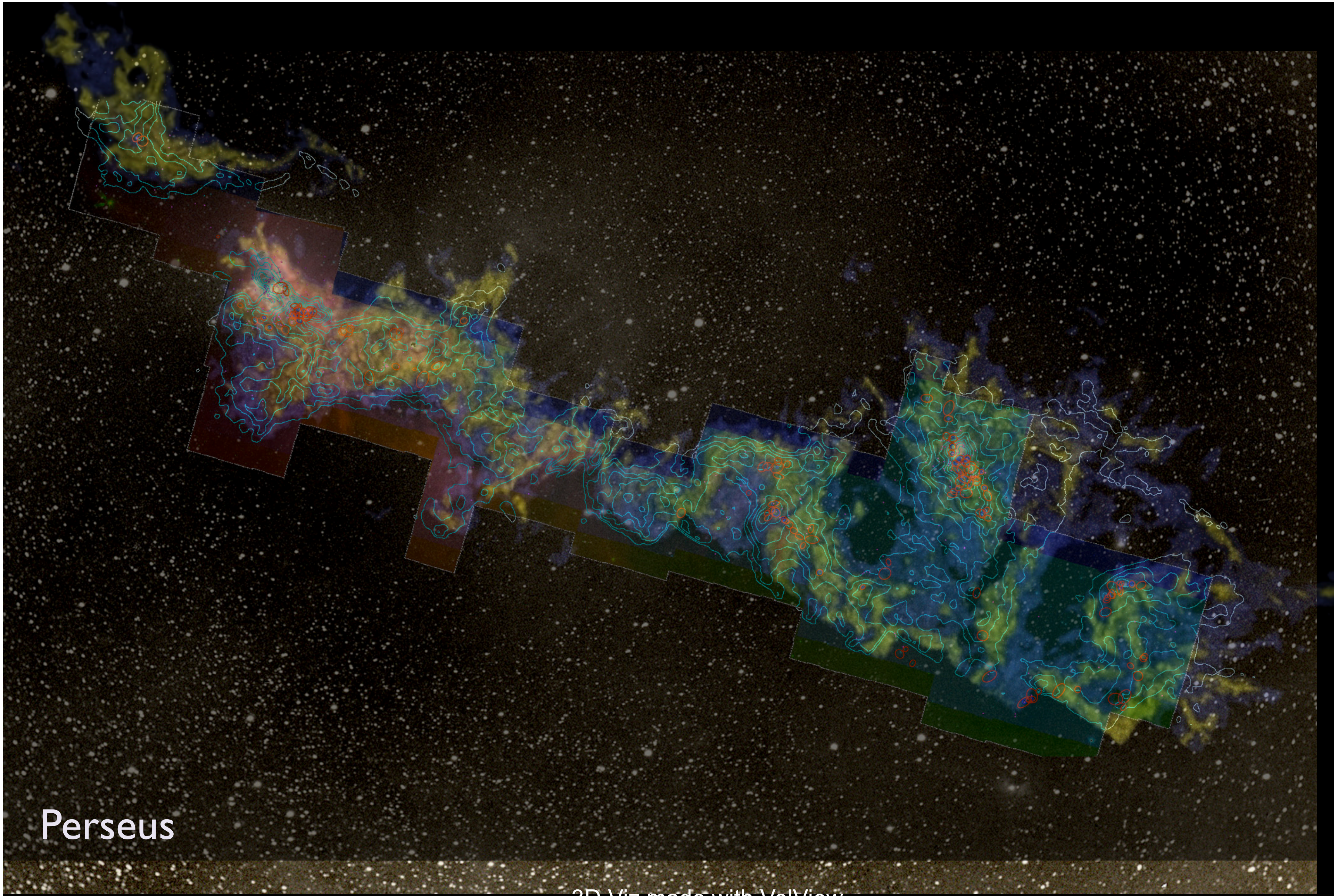
○ sub-mm peak (Hatchell et al. 2005, Kirk et al. 2006)

^{13}CO (Ridge et al. 2006)

▨ mid-IR IRAC composite from c2d data (Foster, Laakso, Ridge, et al. in prep.)

▨ Optical image (Barnard 1927)





3D Viz made with VolView

AstroMed@IIC

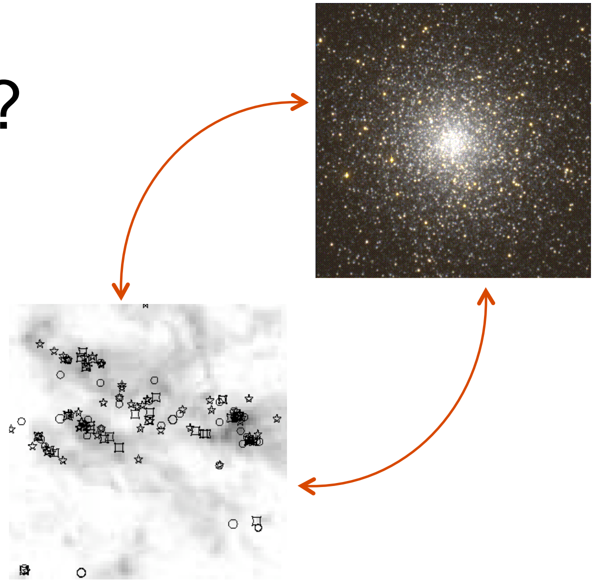
COMPLETE

WIE?

Sternentstehung

- *Wie* bilden sich Sterne?
- Was bestimmt *wo* und *wann* Sterne entstehen?
- Welche physikalische Prozesse *initiieren* und *regulieren* Sternentstehung und ihre *Effizienz*?
- Wie beeinflussen *globale* Eigenschaften der Galaxie die Sternentstehung auf *lokalen* Skalen, und umgekehrt?
- Gibt es verschiedene *Modi* der *SE*? (*isolierte* SE ↔ SE im *Haufen*)

→ **Identifikation der relevanten physikalischen Prozesse!**



47 Tucanae (NAOA)

Taurus (Hartmann 2002)

Gravoturbulente Sternentstehung

- These:

Sternentstehung wird kontrolliert durch das Wechselspiel von Gravitation und Überschallturbulenz!

- Duale Rolle der Turbulenz:
  - *Stabilität auf großen Skalen*
  - *Initiiert Kollaps auf kleinen Skalen*

(Mac Low & Klessen, 2004, Rev. Mod. Phys., 76, 125-194)

Im Detail...

Gravoturbulente Sternentstehung

- *Überschallturbulenz* in der galaktischen Scheibe produziert starke Dichtefluktuationen (in Schocks: $\delta\rho/\rho \propto M^2$)
  - chemischer Phasenübergang: atomar → molekular
  - Kühlungsinstabilität
  - gravitative Instabilität
- Kalte *Molekülwolken* entstehen in Gebieten hoher Dichte
- *Turbulenz* erzeugt Dichtestruktur, *Gravitation* selektiert für Kollaps

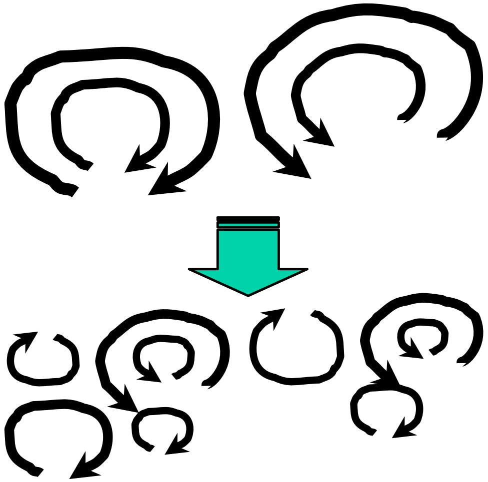
$\longrightarrow$ ***GRAVOTUBULENTE FRAGMENTATION***

- *Turbulente Kaskade:* Lokale Kompression *im Inneren* der Wolke führt zu Kollaps → *Sterne* und *Sternhaufen*

(Mac Low & Klessen, 2004, Rev. Mod. Phys., 76, 125-194)

Eigenschaften der Turbulenz

- Laminarer Fluss wird *turbulent* bei *großen* *Reynolds* Zahlen



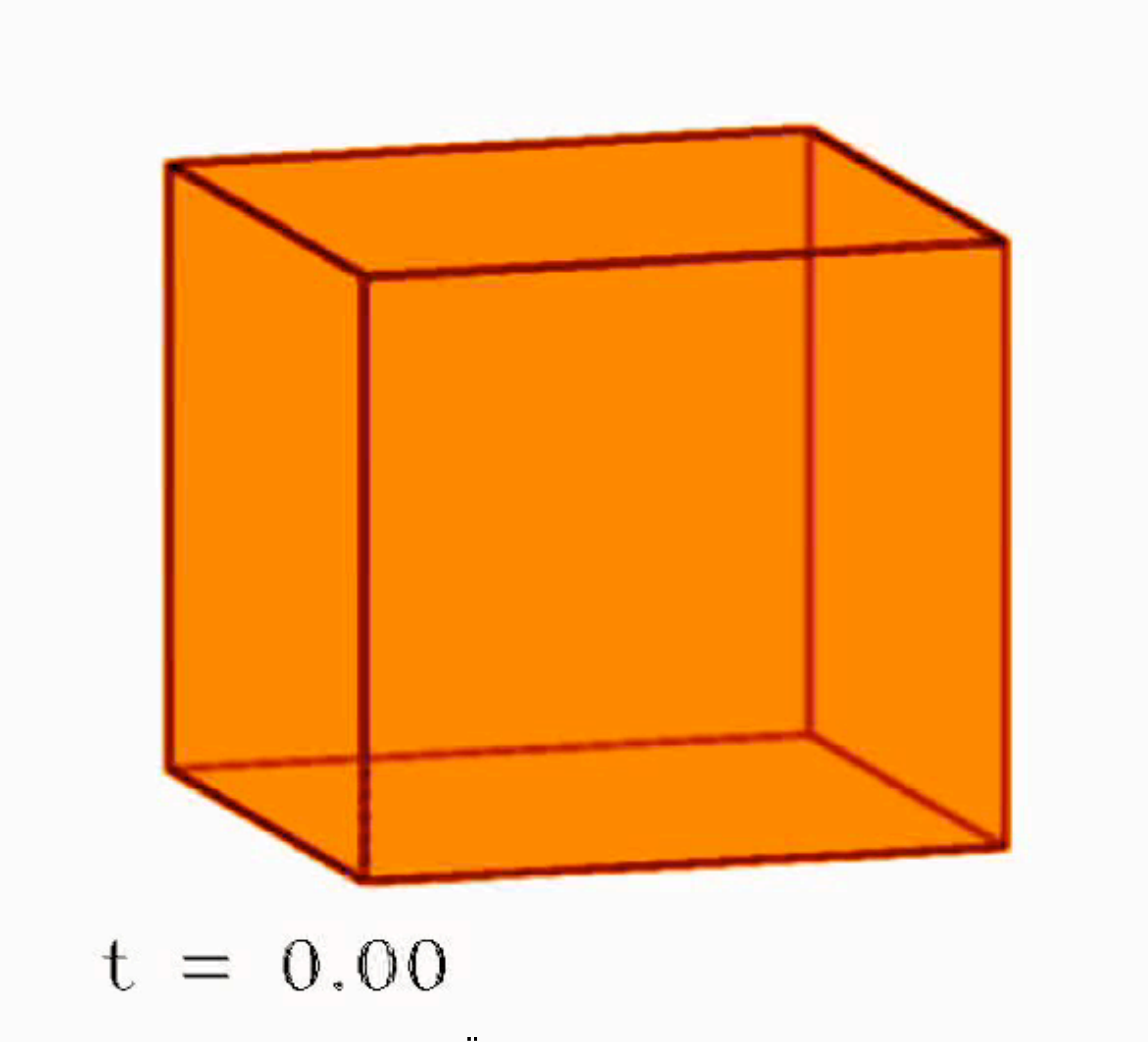
$$\boxed{\mathrm{Re} = \frac{\text{Advektion}}{\text{Dissipation}} = \frac{VL}{\nu}}$$

V= typical velocity on scale L, ν = viscosity, Re > 1000

- *vortex streching* --> Turbulenz ist intrinsisch *anisotrop!*

  (nur auf großen Skalen kann Homogenität & Isotropie gelten, in statistischem Sinn; z.B. Landau & Lifschitz, Chandrasekhar, Taylor, etc.)

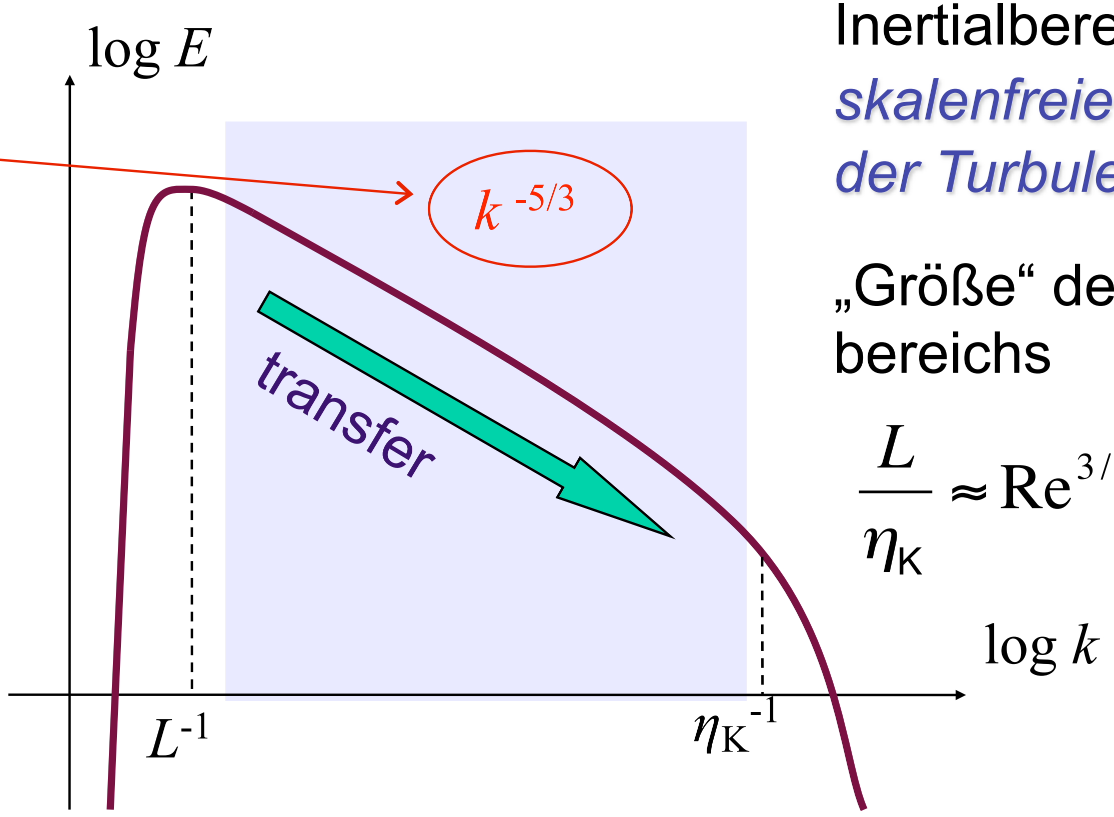
  (ISM Turbulenz: Stöße & B-Feld erzeugen weitere Inhomogenitäten)



rein hydrodynamische Überschallturbulenz in 3D
(mit Eigengravitation bei t ≥ 4.00) (Klessen et al. 2000)

Turbulente Kaskade

Kolmogorov (1941) Theorie der inkompressiblen Turbulenz



Inertialbereich:
skalenfreies Verhalten der Turbulenz

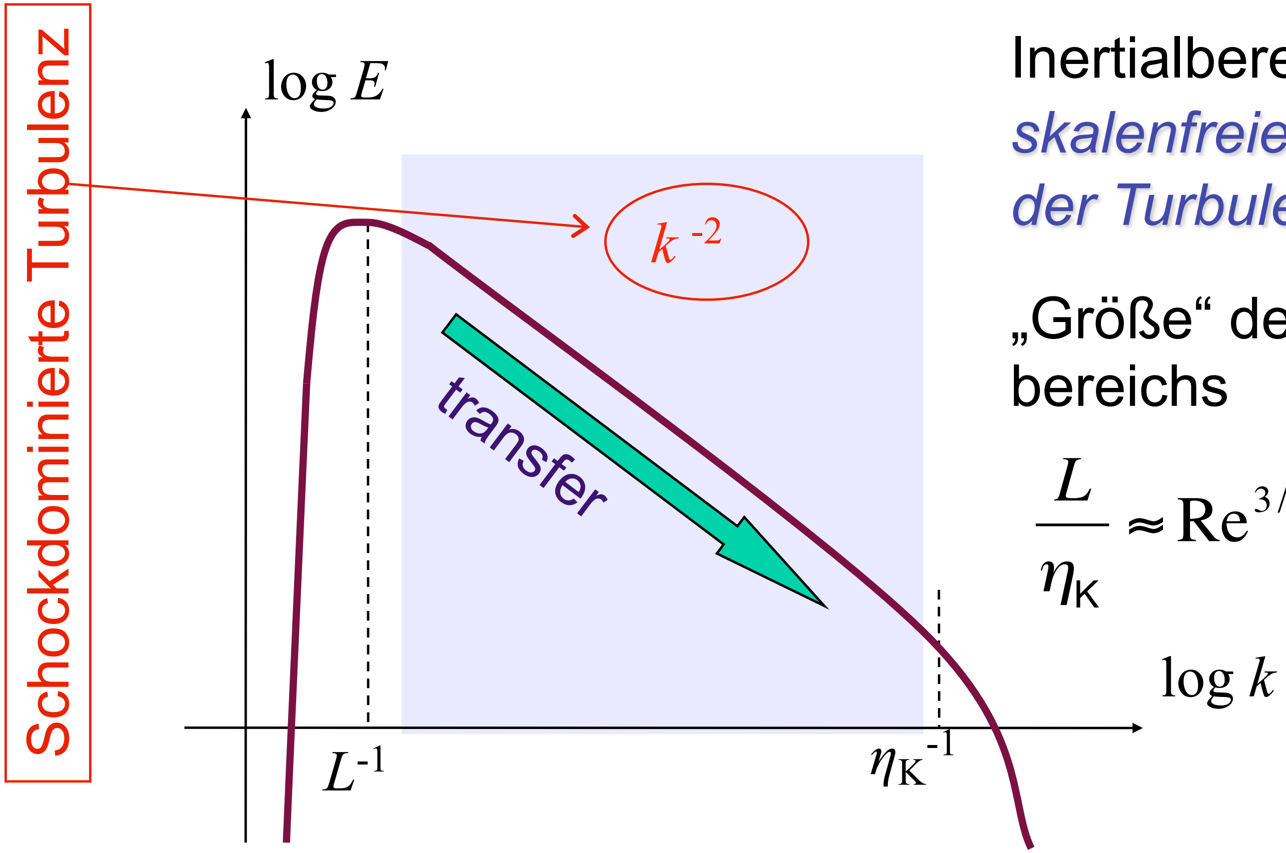
„Größe“ des Intertialbereichs

$$\frac{L}{\eta_K} \approx \mathrm{Re}^{3/4}$$

Skala des Energieeintrages

Skala der Energiedissipation

Turbulente Kaskade



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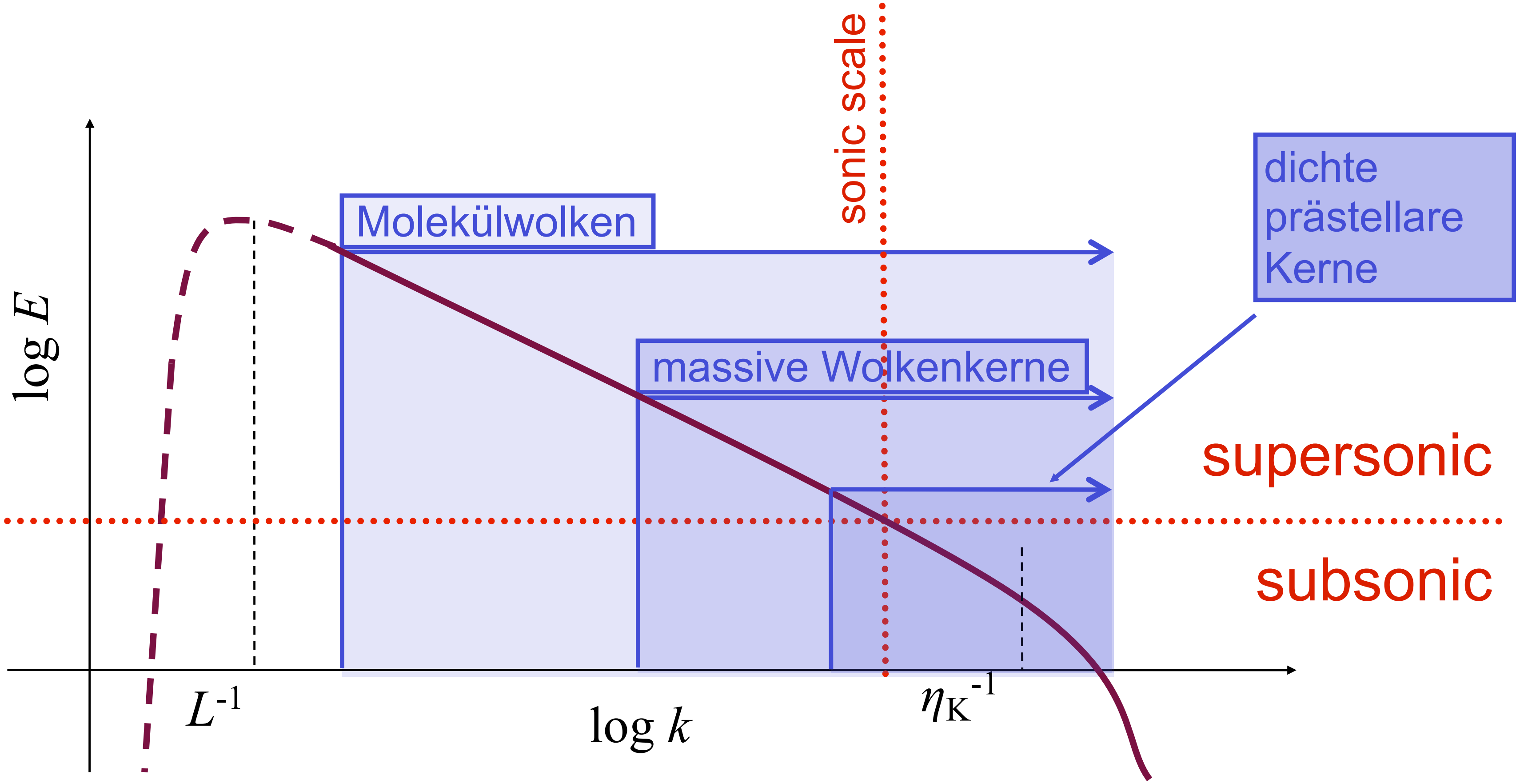
Skala der Energiedissipation

Turbulente Kaskade in der ISM



Energiequelle & -skala ist UNBEKANNT (Supernovae, Winde, Dichtewellen ?)

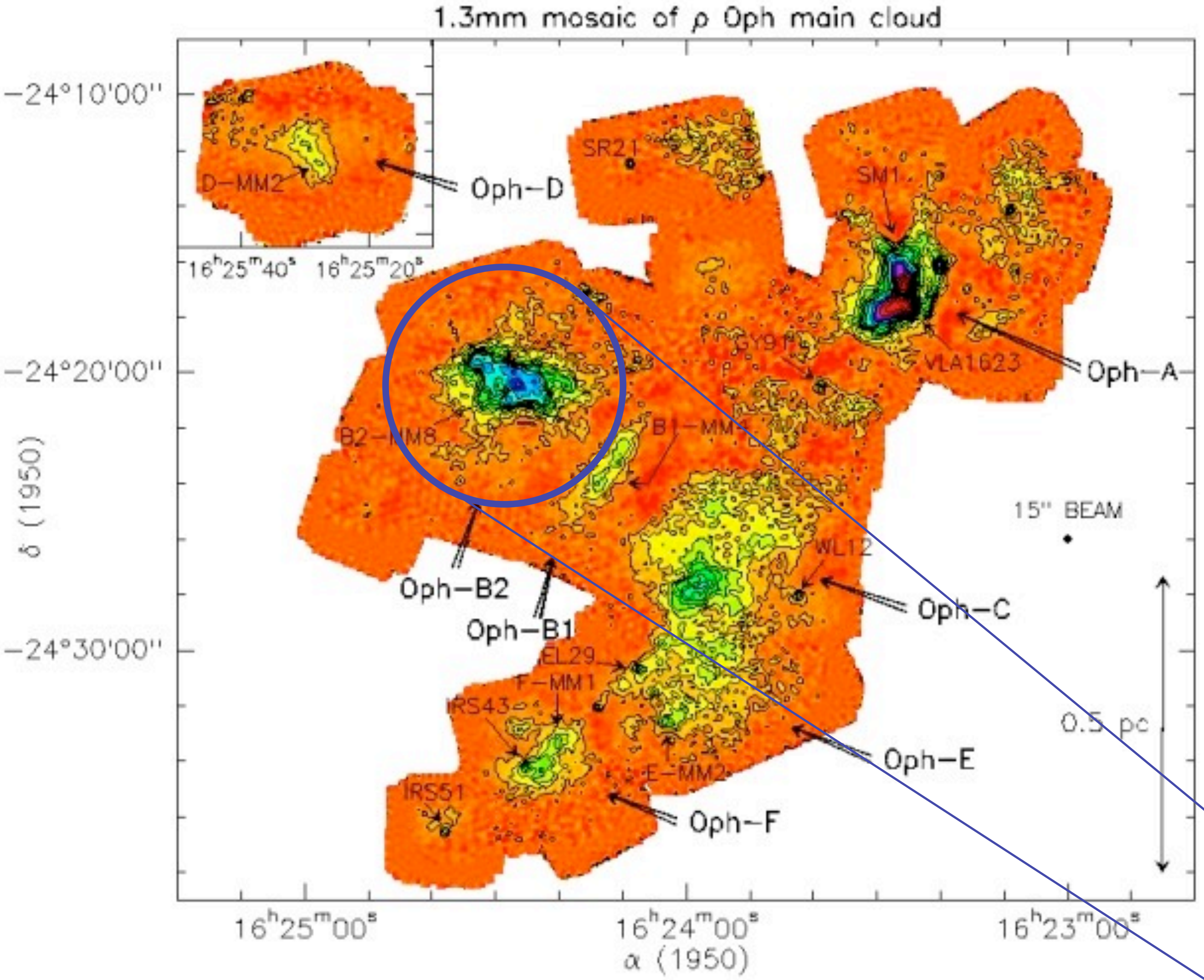
$\sigma_{rms} << 1$ km/s

$\mathcal{M}_{rms} \leq 1$

$L \approx 0.1$ pc

Dissipationsskala auch unbekannt (ambipolare Diffusion, molekulare Diffusion?)

Dichtestruktur von MW



(Motte, André, & Neri 1998)

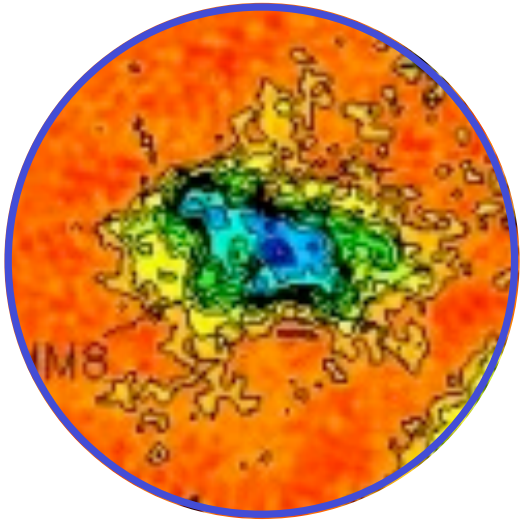
Molekülwolken sind hochgradig inhomogen.

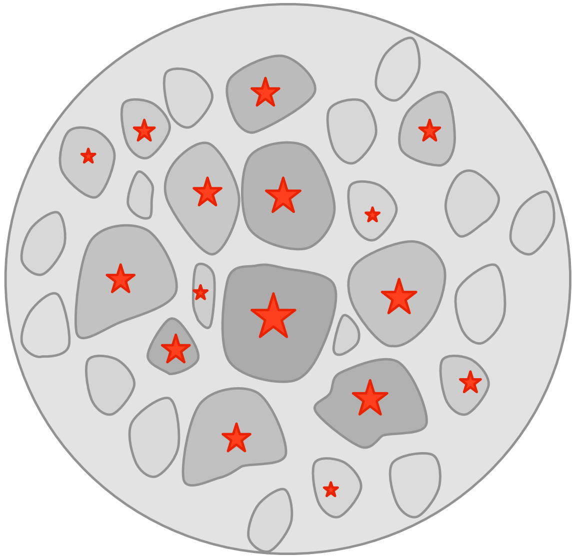
Sterne bilden sich in den dichtesten und kältesten Teilen der Wolke.

ρ-Ophiuchus in Staubemission

genauere Betrachtung

Entwicklung von Wolkenkernen





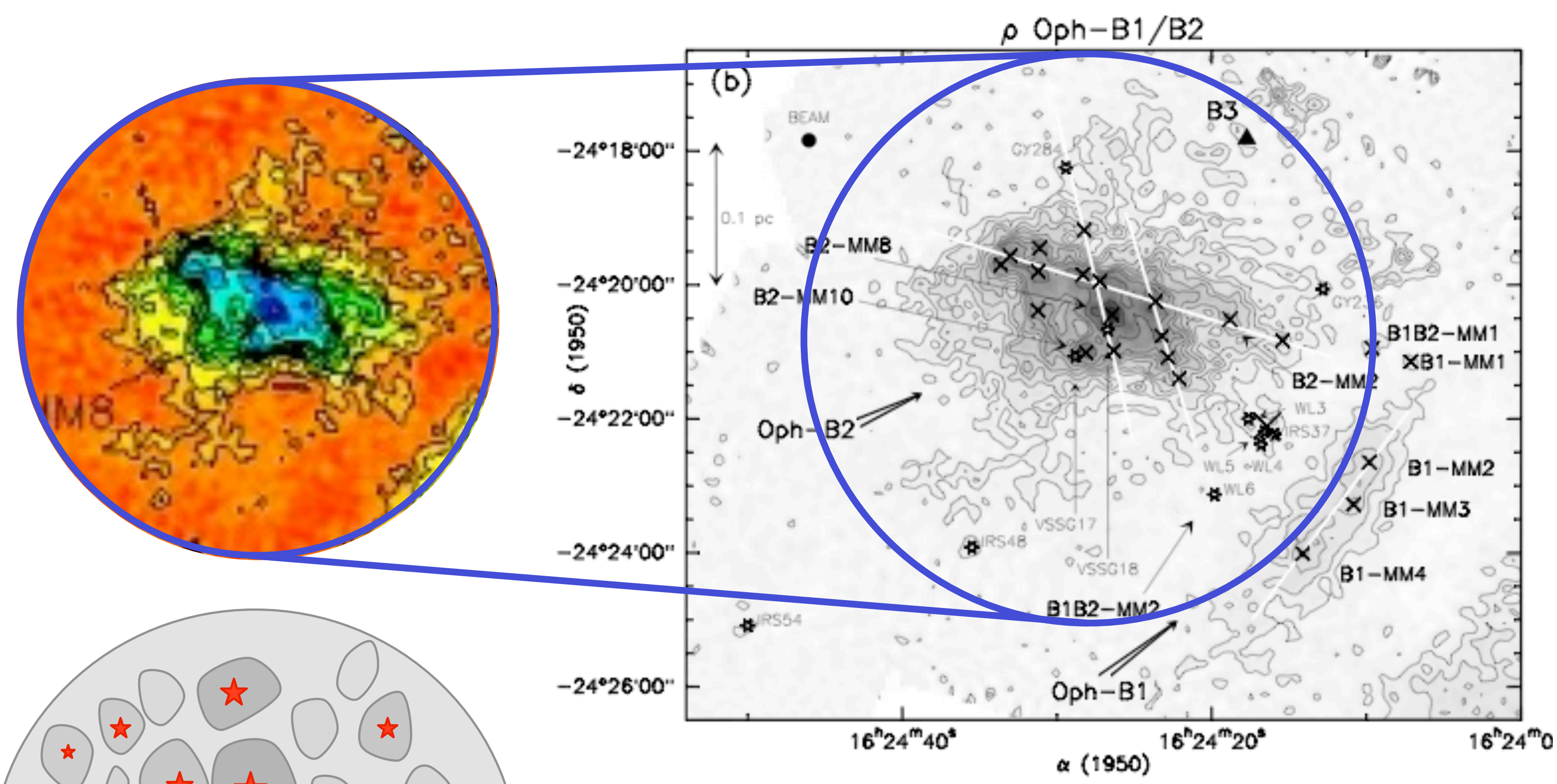
- *FRAGE*:
  Bildet sich ein einzelner massereicher Stern, oder ein Sternhaufen mit massearmen Sternen?

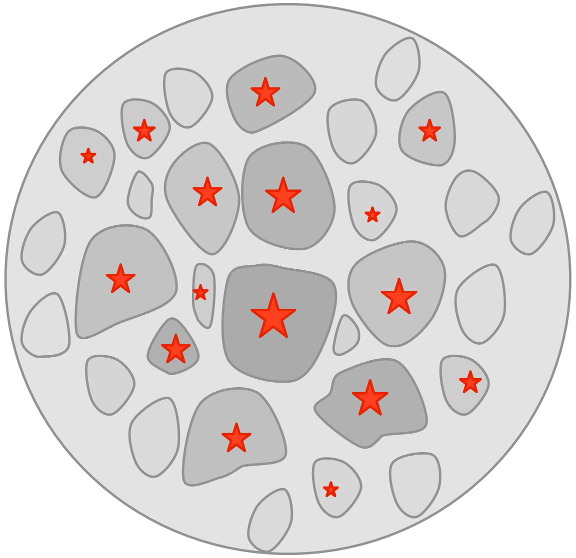
- Turbulente Kaskade „geht durch“ Wolkenkern
  --> KEINE *Skalenseparation* möglich
  --> KEINE *effektive Schallgeschwindigkeit*
- Turbulenz ist überschallschnell!
  --> Erzeugt große Dichtekontraste:

  $$\delta\rho/\rho \approx M^2$$

  --> mit typischerweise $M \approx 10$ --> $\delta\rho/\rho \approx 100$!
- Viele schock-induzierte Dichtefluktuationen sind gravitativ instabil und kollabieren.
- --> Vorhersage: *Kern fragmentiert und bildet Sternhaufen.*

Entwicklung von Wolkenkernen



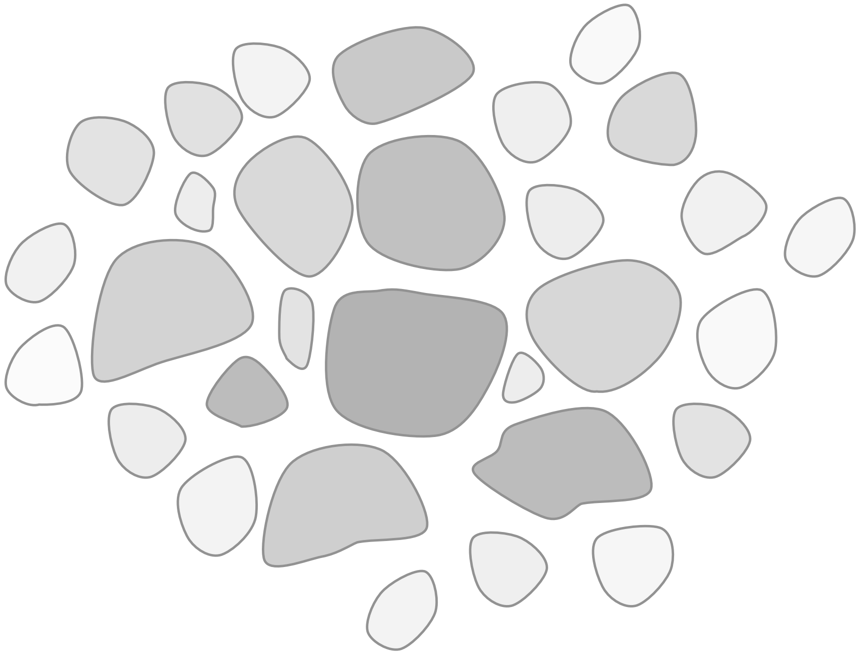


Beispiel ρ-Oph B1/B2: Kern enthält mehrere protostellare Objekte

(“starless” cores are denoted by ×, cores with embedded protostars by ☆)

Entwicklung von Wolkenkernen

Was passiert bei mehreren (wechselwirkenden) Kernen?



Zwei Extremfälle:

(1) **Turbulenz dominiert Energiebudget:**

$\alpha = E_{kin}/|E_{pot}| > 1$

--> keine Wechselwirkung der Kerne untereinander

--> *Stellarer Massenzuwachs* bestimmt durch *isolierten Kollaps*

--> *loser Haufen massearmer Sterne*

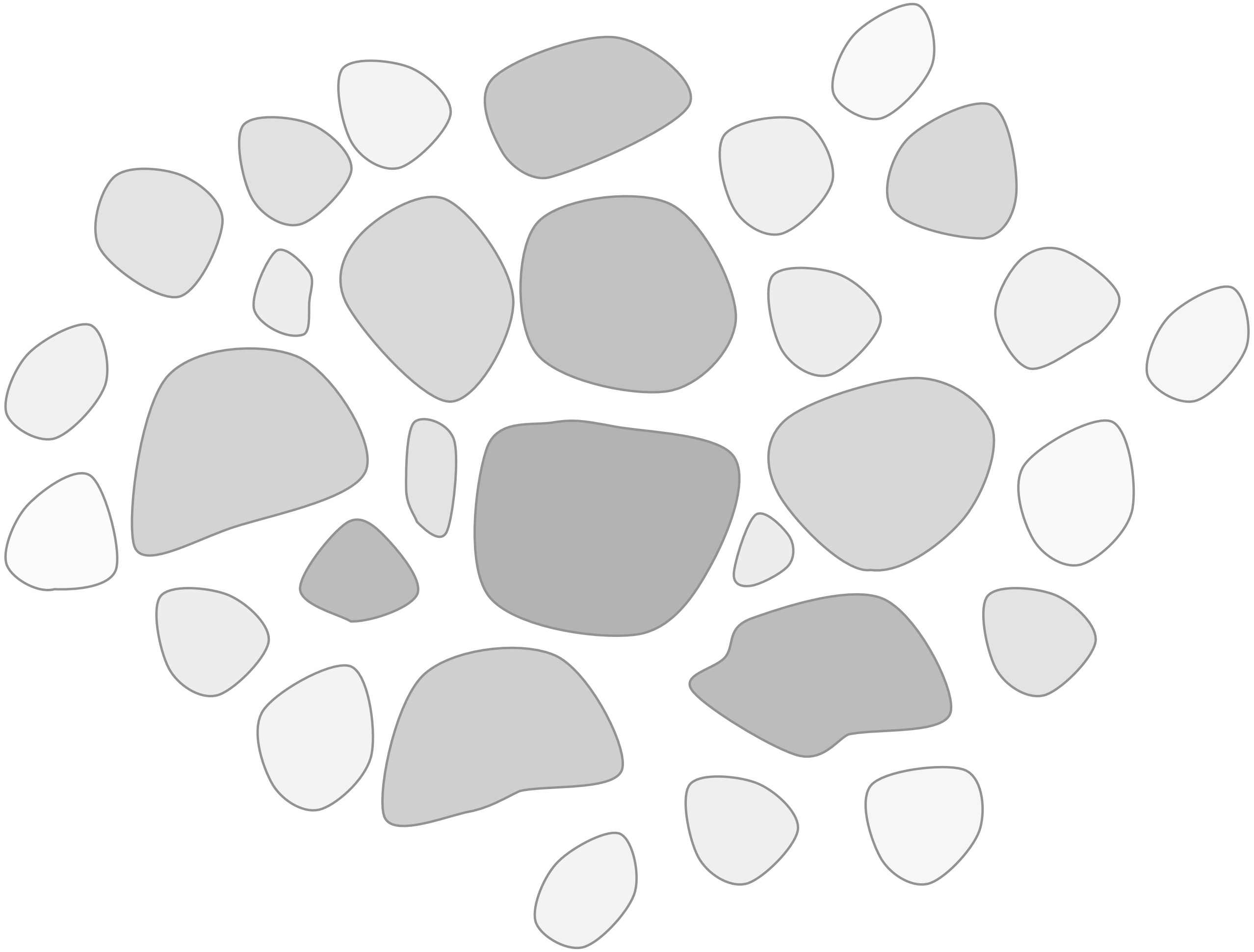
(2) **Turbulenz zerfällt**, dh. **Gravitation dominiert:** $\alpha = E_{kin}/|E_{pot}| < 1$

--> *globale Kontraktion*

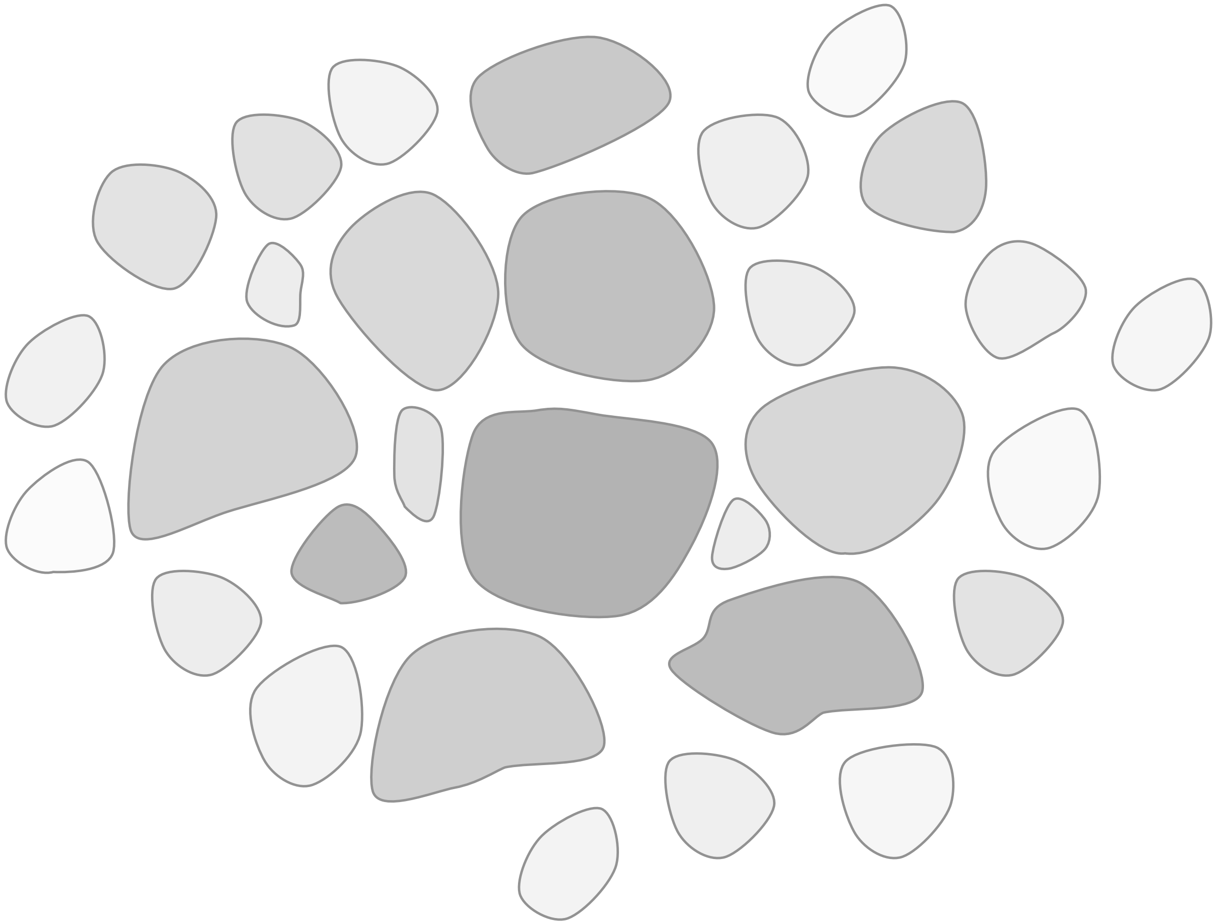
--> Kerne *wechselwirken* während sie kollabieren

--> *Kompetitives Wachstum*

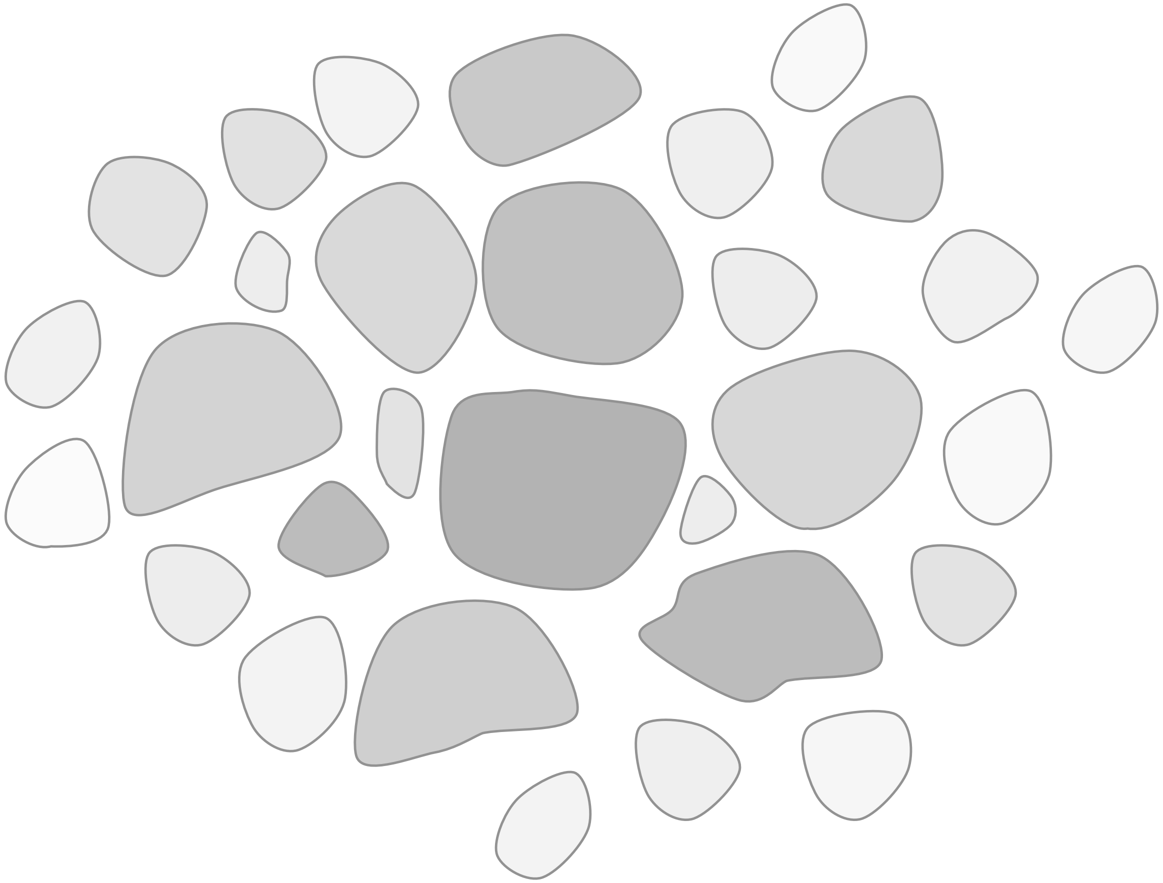
--> *dichter Haufen, enthält massereiche Sterne*



Turbulenz erzeugt Hierarchie von Klumpen



Turbulenz dissipiert, Kontraktion setzt ein



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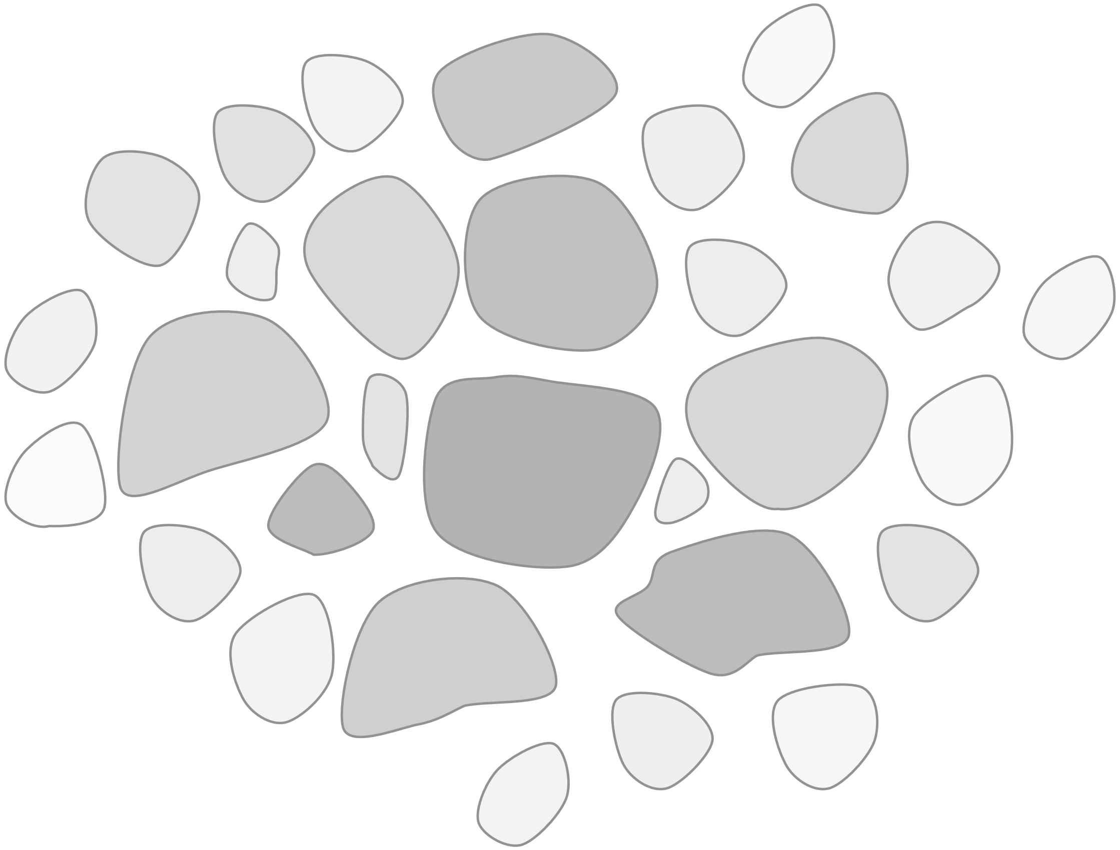
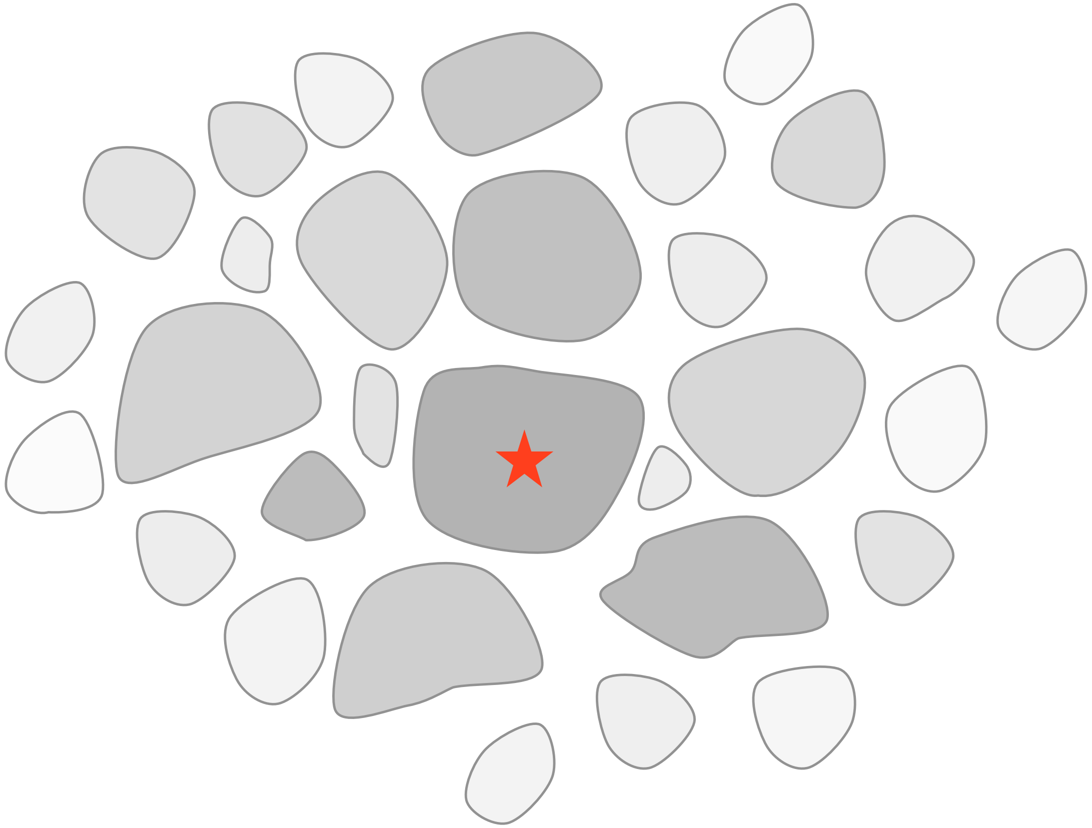


während Region kontrahiert können einzelne Klumpen kollabieren und Sterne bilden



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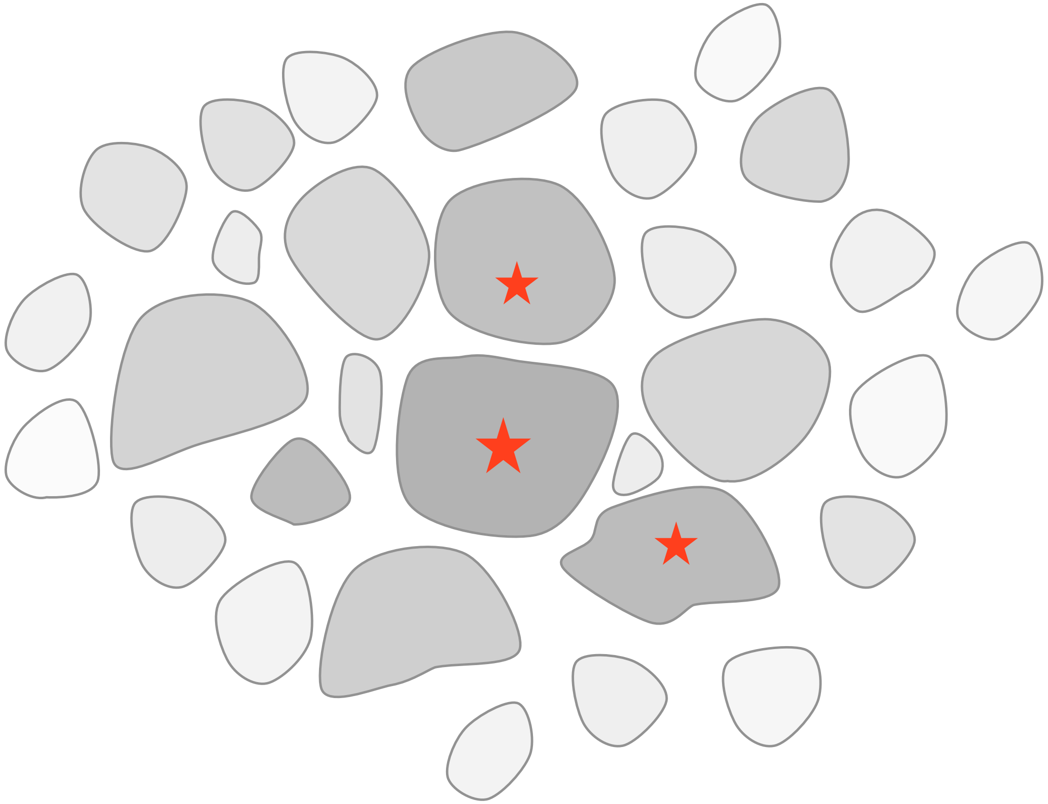
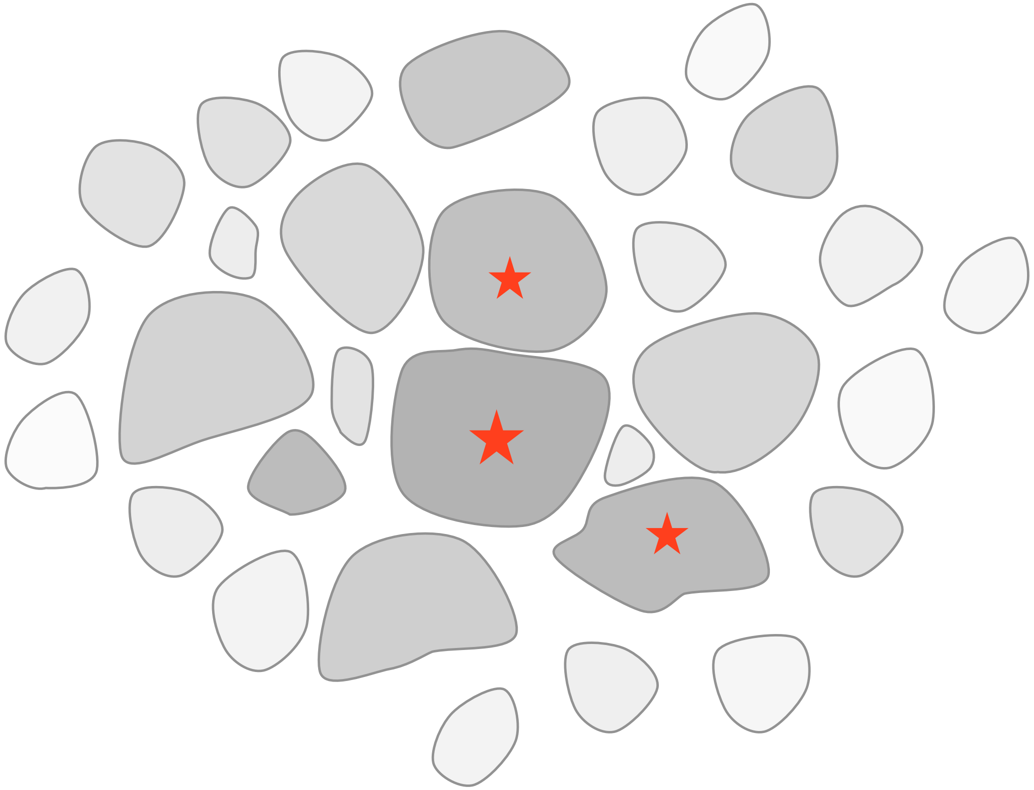


einzelne Klumpen kollabieren und bilden Sterne



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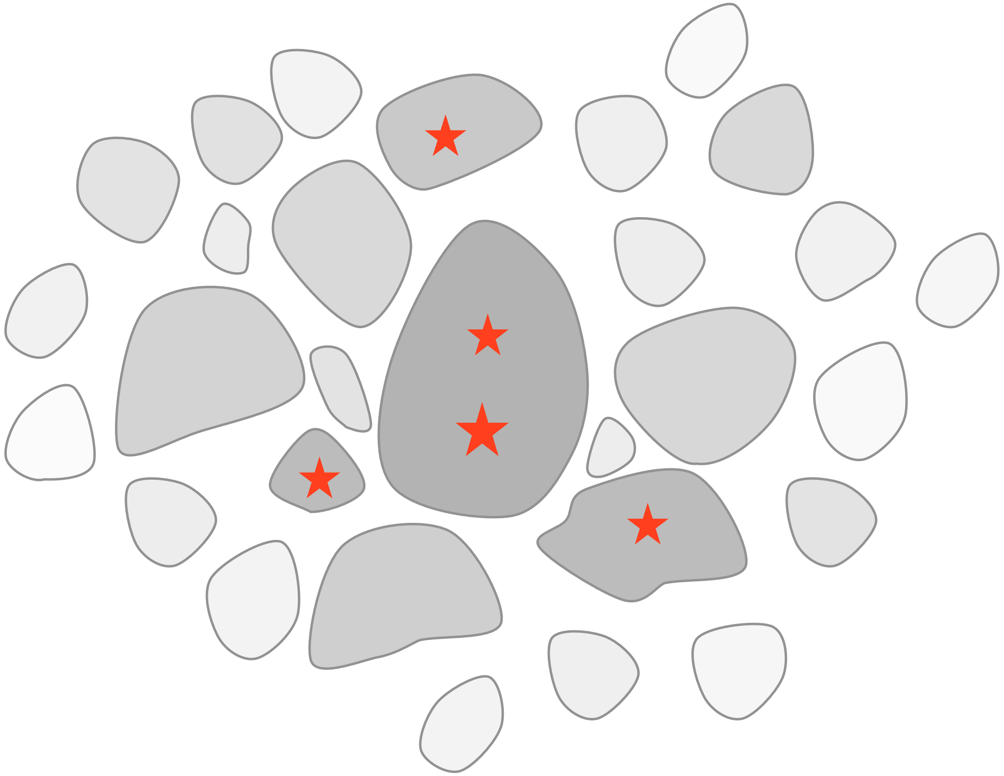
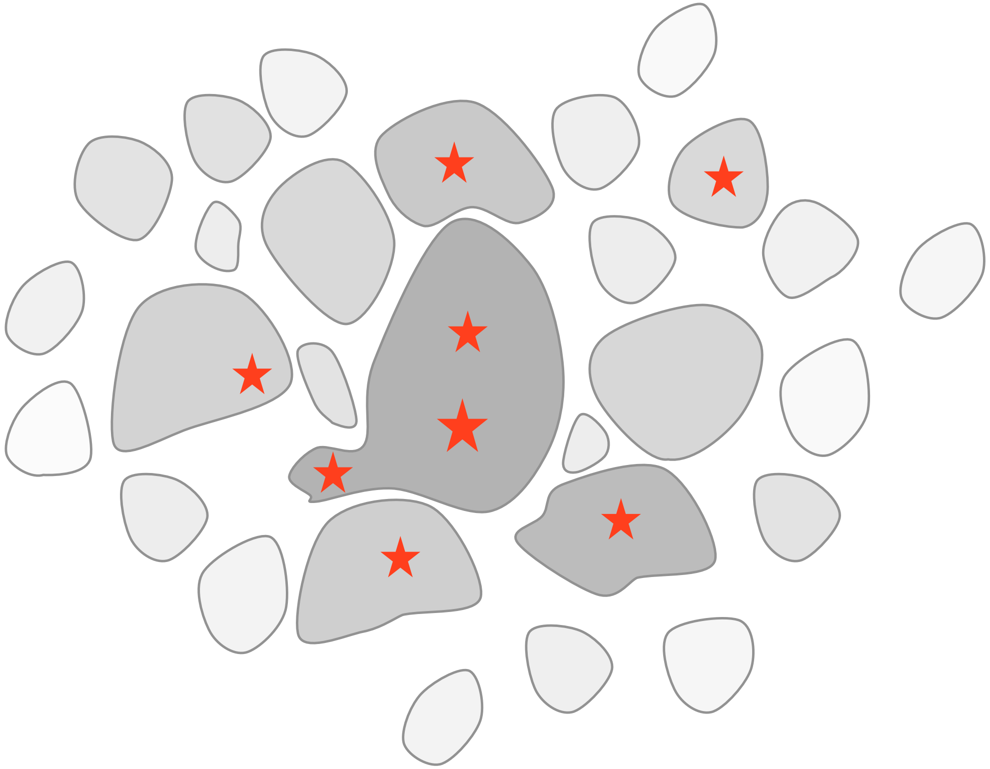


$\alpha = E_{kin}/|E_{pot}| < 1$

Im *dichten Haufen* können Klumpen verschmelzen während sie kollabieren --> sie enthalten nun mehrere Protosterne



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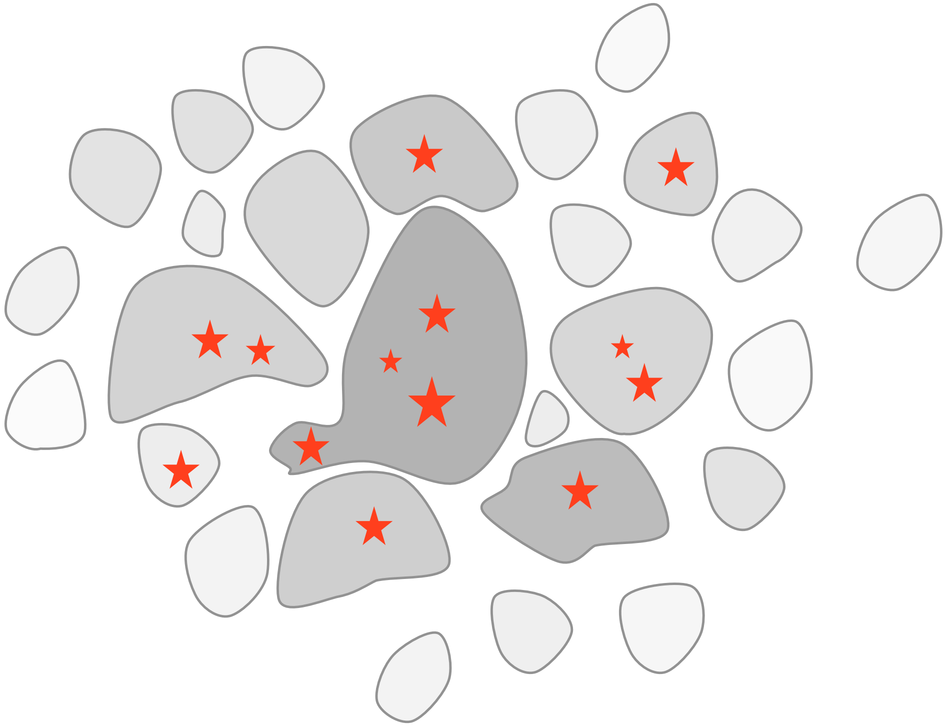

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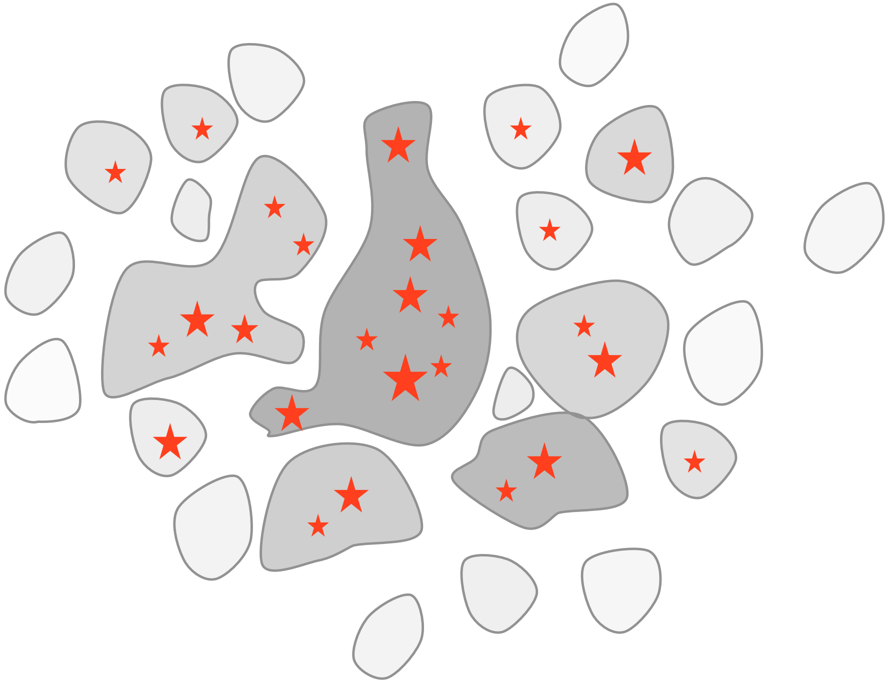
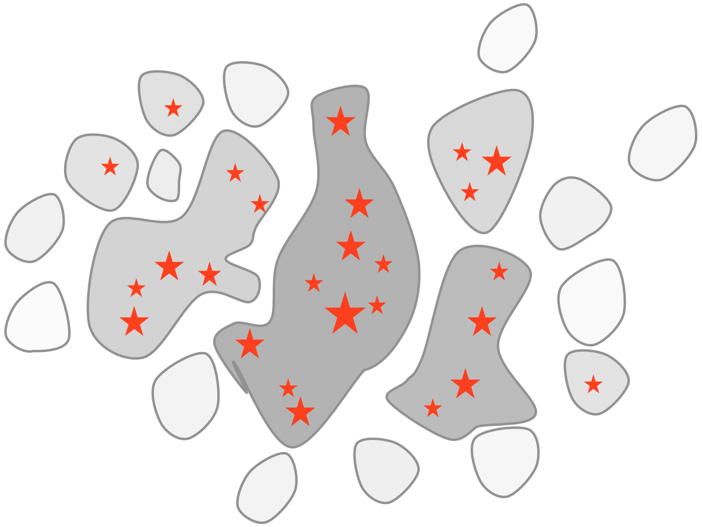


Im *dichten Haufen* wird kompetitives Wachstum wichtig



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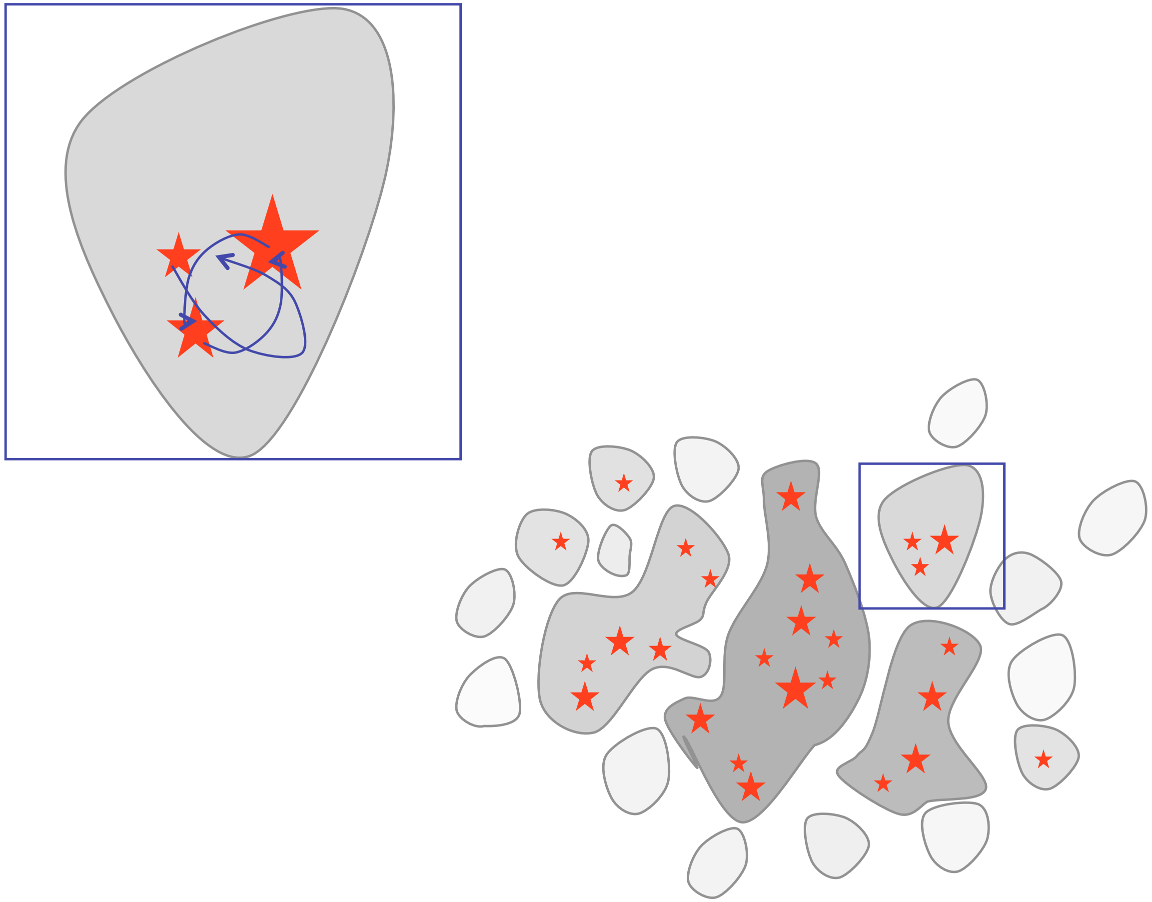
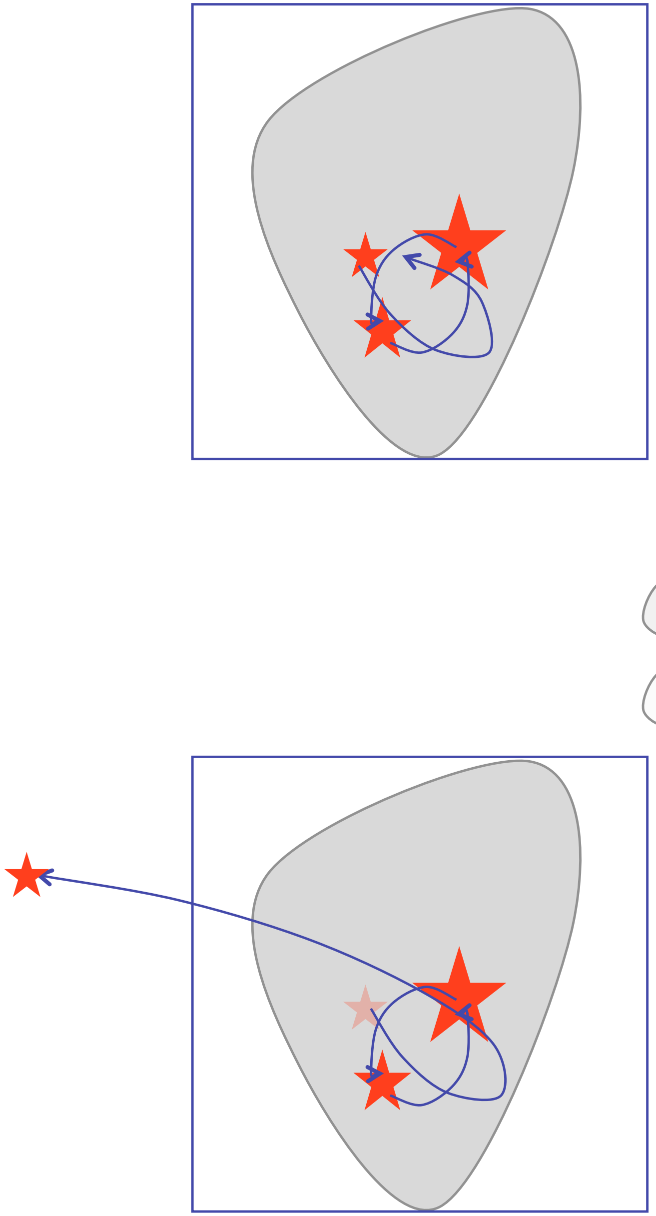
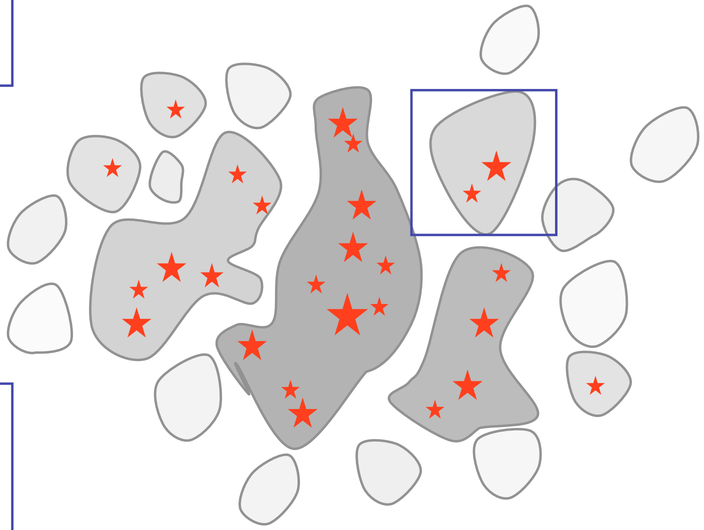
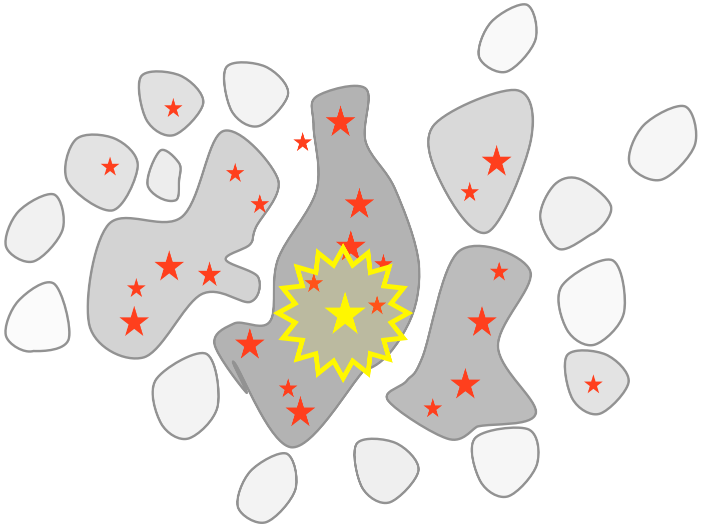


Im *dichten Haufen* beeinflussen stellardynamische Prozesse das Wachstum

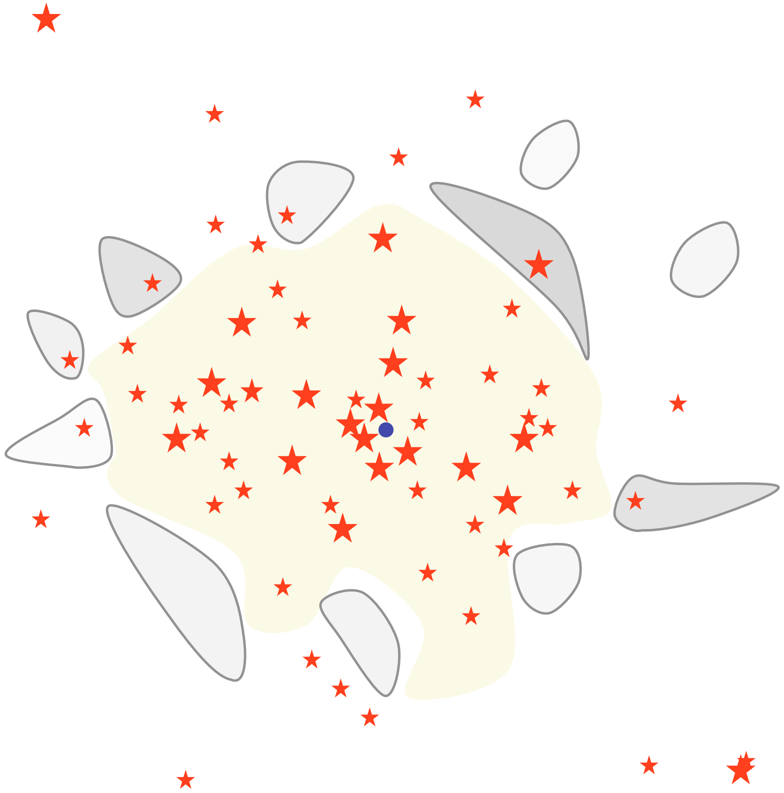




massearme Objekte können herausgeschleudert werden --> Ende der Akkretion



Feedback beendet die Sternbildung



Resultat: *Sternhaufen*, evtl. umgeben von HII-Region

NGC 602 in the LMC: Hubble Heritage Image

Resultat: *Sternhaufen*, evtl. umgeben von HII-Region

Vorhersagen

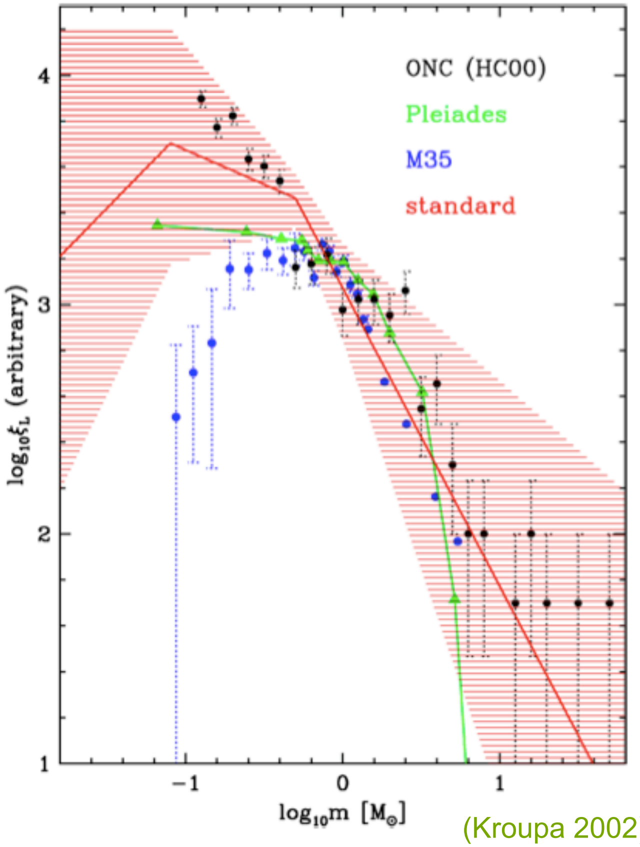
Was können wir lernen?

- *Globale Eigenschaften* *(statistische Betrachtung)*
  - Sternentstehungseffizienz & Zeitskalen
  - Stellare Massenfunktion (IMF)
  - Dynamik junger Sternhaufen
  - Statistische Beschreibung von Überschallturbulenz (PDF's, Δ-*Var.)*
  - Mischungseigenschaften von Überschallturbulenz (wichtig für die chemische Entwicklung der Milchstraße)
- *Lokale Eigenschaften* *(Studium von Einzelobjekten)*
  - Eigenschaften protostellarer Gasklumpen (z.B. Form, radiales Dichteprofil, Geschwindigkeitsfeld)
  - Akkretionsgeschichte einzelner Protosterne (dM/dt vs. t, j vs. t)
  - Doppelsterne (Exzentrizitäten, Massenverhältnisse, usw.)
  - SED's von Protosternen, PMS Entwicklungsreihen

(Mac Low & Klessen, 2004, Larson 2005, Ballesteros-Paredes et al. 2007))

Stellare Massenfunktion

- die anfängliche Massenverteilung scheint einem “universellen” Gesetz zu folgen...



(Kroupa 2002)

Orion, NGC 3603, 30 Doradus
(Zinnecker & Yorke 2007)

IMF

- distribution of stellar masses depends on
  - turbulent initial conditions
    --> mass spectrum of prestellar cloud cores
  - collapse and interaction of prestellar cores
    --> competitive accretion and *N*-body effects
  - thermodynamic properties of gas
    --> balance between heating and cooling
    --> EOS (determines which cores go into collapse)
  - (proto) stellar feedback terminates star formation
    ionizing radiation, bipolar outflows, winds, SN

(e.g. Larson 2003, Prog. Rep. Phys.; Mac Low & Klessen, 2004, Rev. Mod. Phys, 76, 125 - 194)

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example: model of Orion cloud

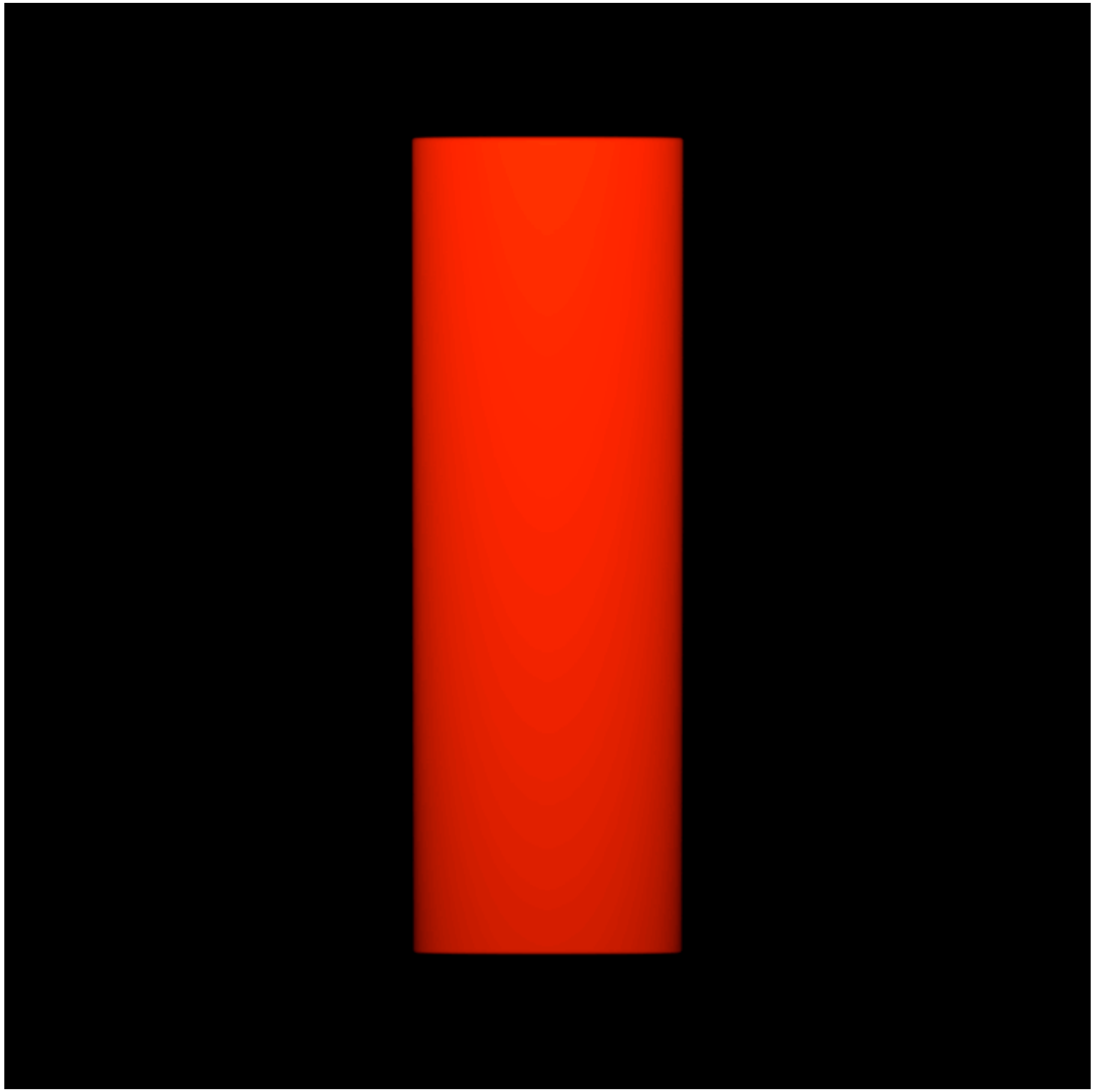
„model“ of Orion cloud: 15.000.000 SPH particles, 10^4 M_{sun} in 10 pc, mass resolution 0,02 M_{sun}, forms ~2.500 „stars“ (sink particles)

isothermal EOS, top bound, bottom unbound

has clustered as well as distributed „star“ formation

efficiency varies from 1% to 20%

develops full IMF
(distribution of sink particle masses)



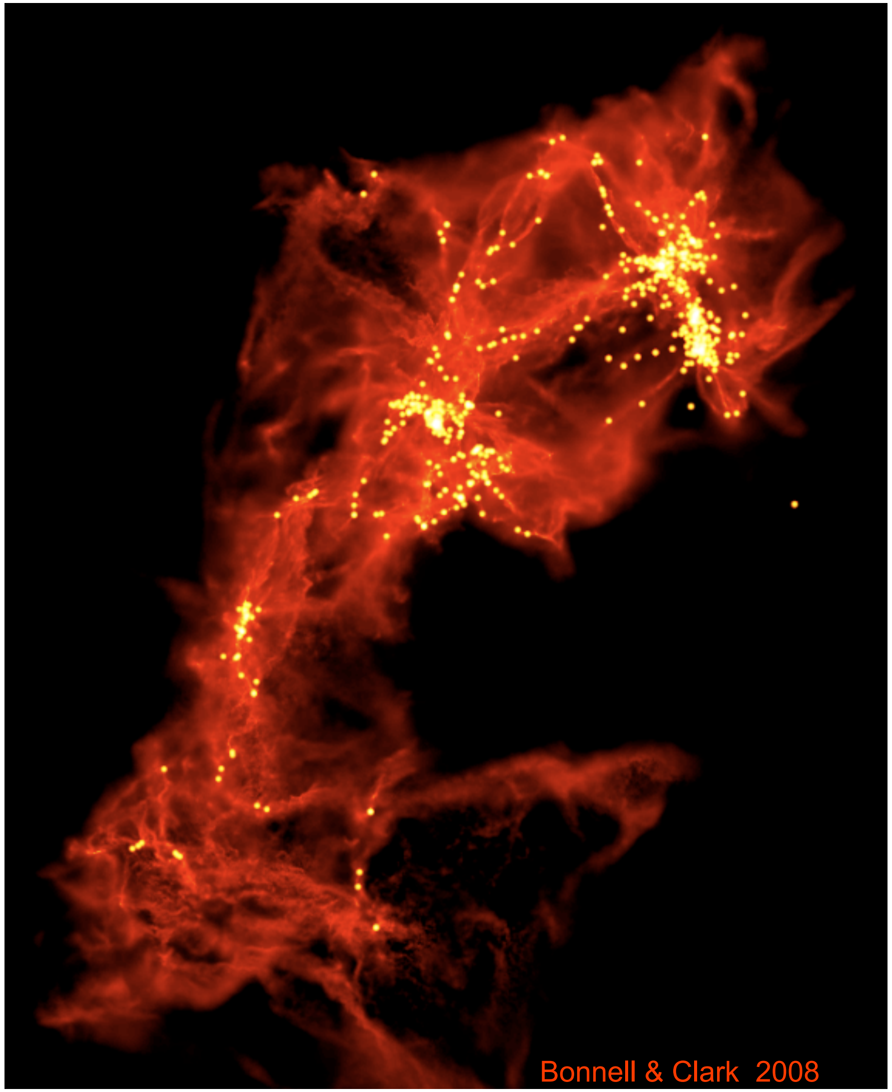
(Bonnell & Clark 2008)

example: model of Orion cloud

„model“ of Orion cloud: 15.000.000 SPH particles, 10^4 M_{sun} in 10 pc, mass resolution 0,02 M_{sun}, forms ~2.500 „stars“ (sink particles)

MASSIVE STARS
- form early in high-density gas clumps (cluster center)
- high accretion rates, maintained for a long time

LOW-MASS STARS
- form later as gas falls into potential well
- high relative velocities
- little subsequent accretion



Bonnell & Clark 2008

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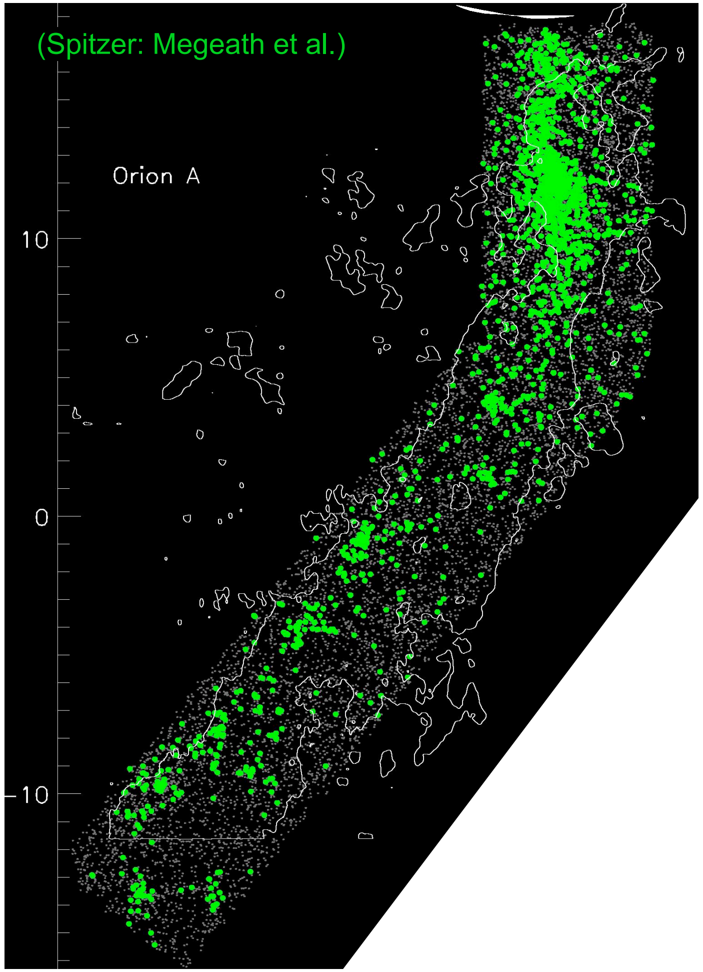
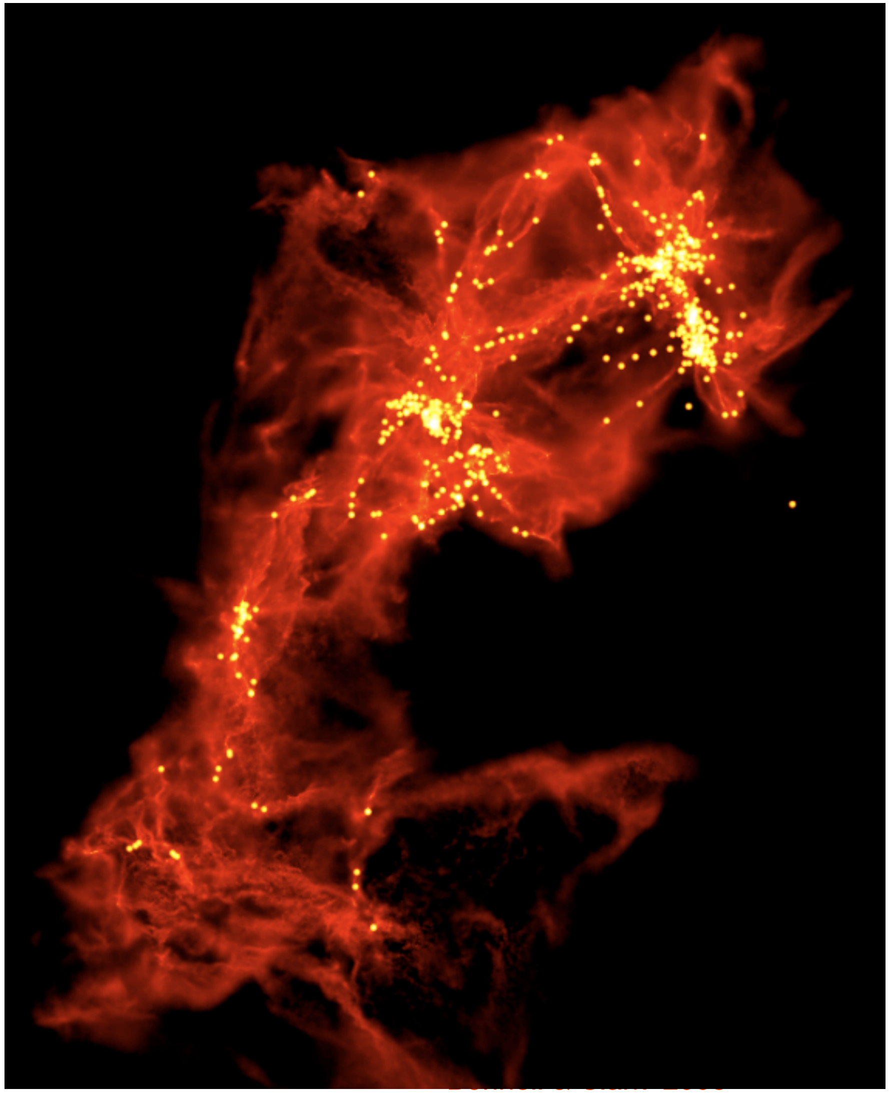
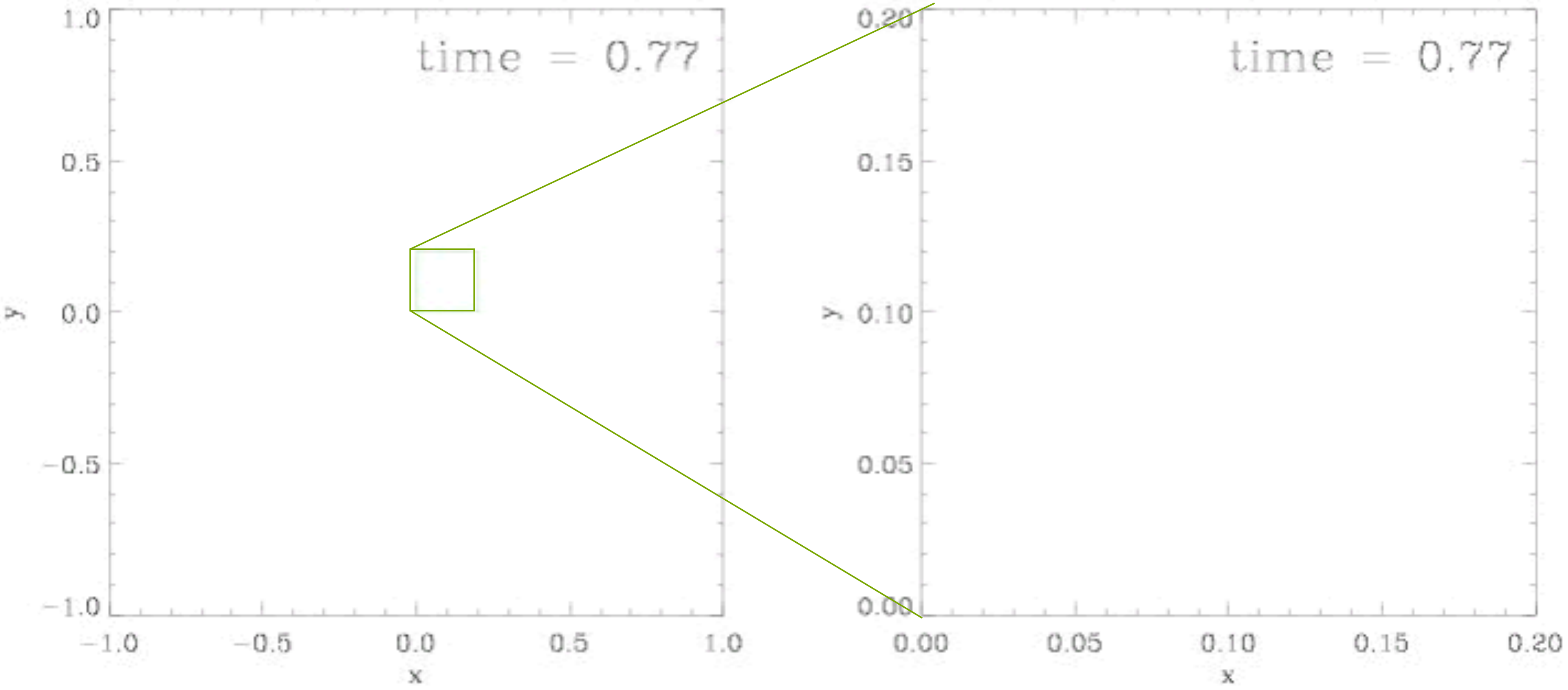




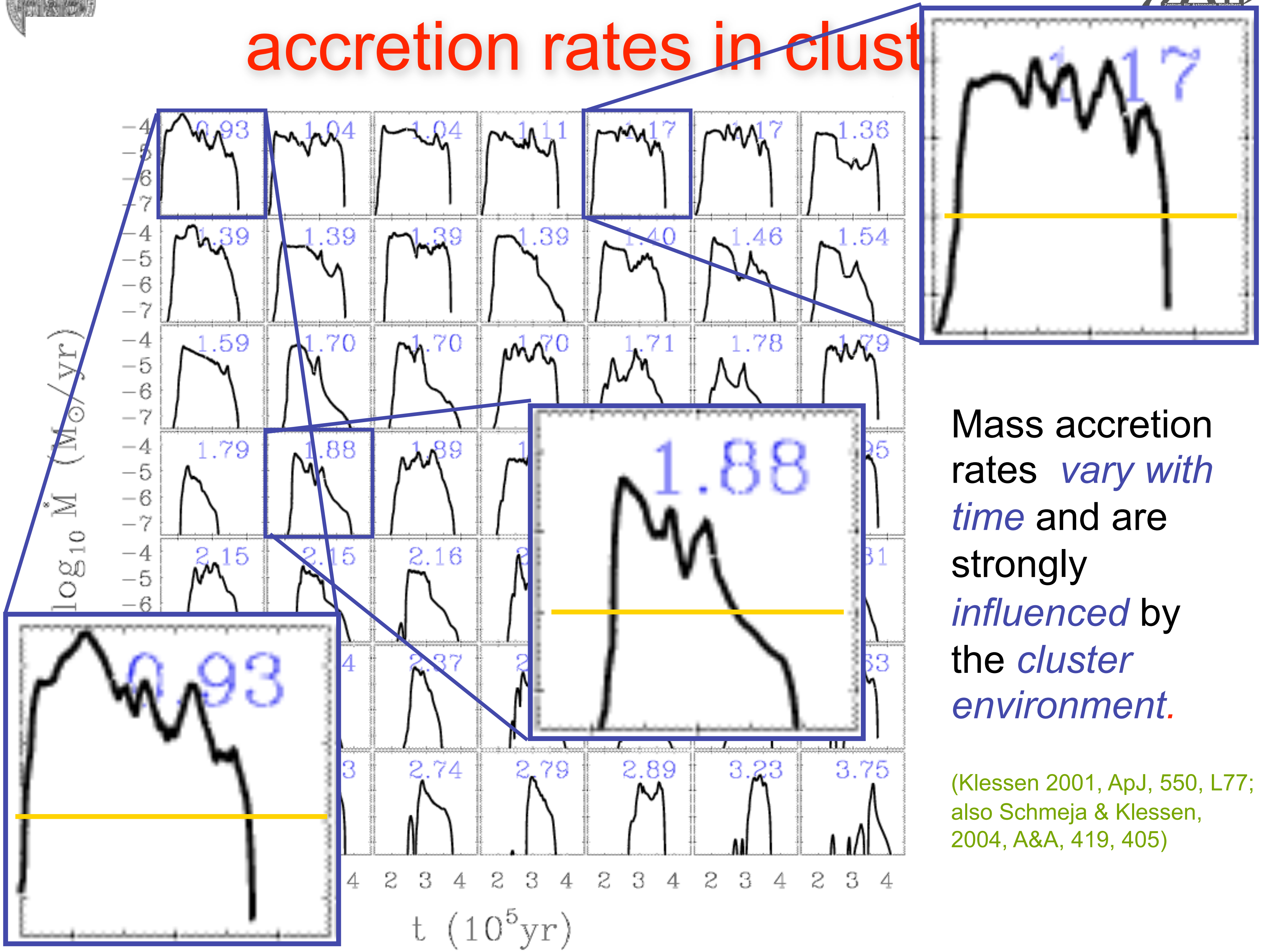
dynamics of nascent star cluster

in dense clusters protostellar interaction may be come important!



Trajectories of protostars in a nascent dense cluster created by gravoturbulent fragmentation
(from Klessen & Burkert 2000, ApJS, 128, 287)

accretion rates in clust



Mass accretion rates *vary with time* and are strongly *influenced* by the *cluster environment*.

(Klessen 2001, ApJ, 550, L77; also Schmeja & Klessen, 2004, A&A, 419, 405)

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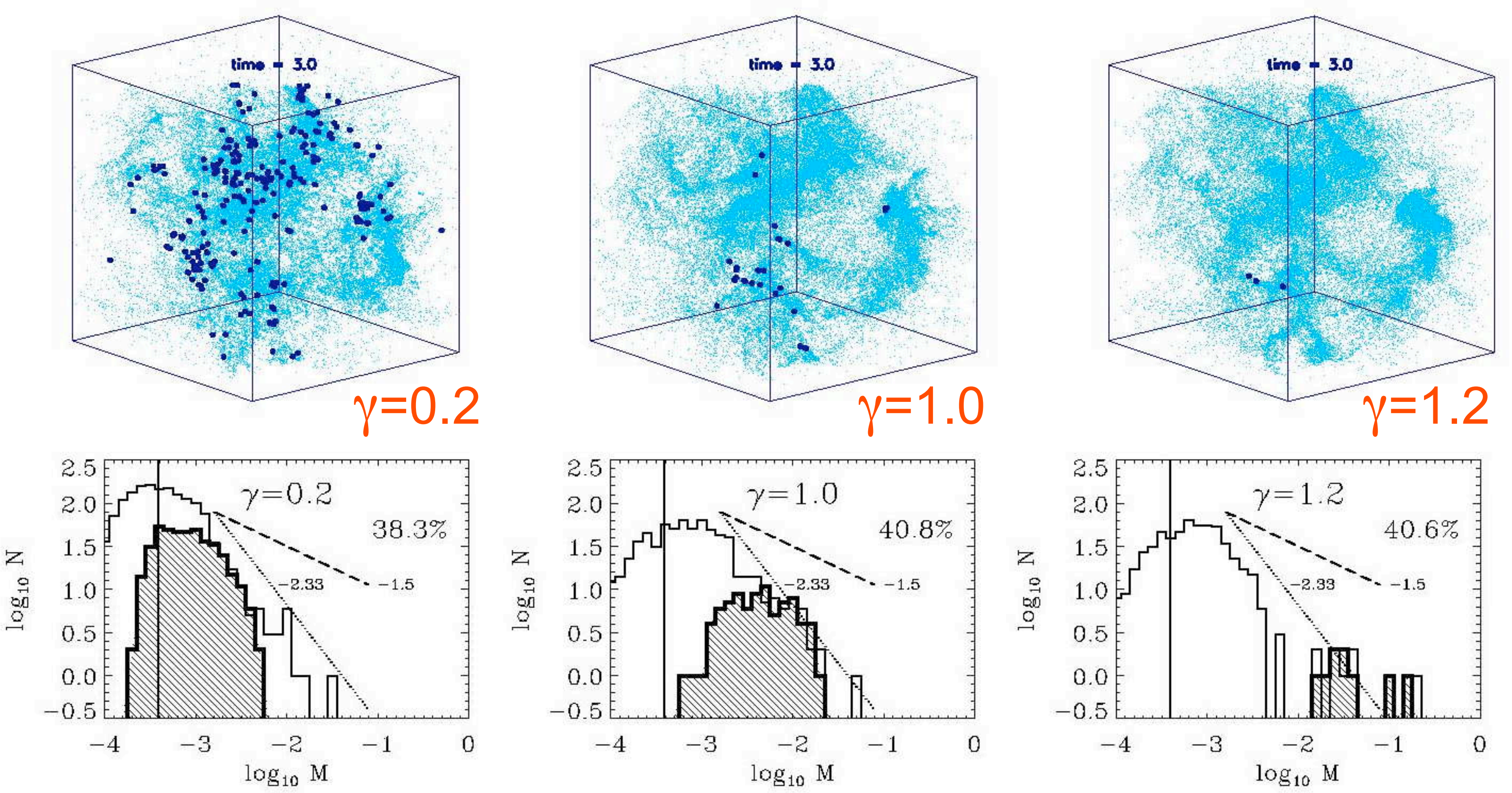
dependency on EOS

- degree of fragmentation depends on *EOS!*

- polytropic EOS: $\mathbf{p} \propto \rho^{\gamma}$
- $\gamma<1$: dense cluster of low-mass stars
- $\gamma>1$: isolated high-mass stars

(see Li, Klessen, & Mac Low 2003, ApJ, 592, 975; also Kawachi & Hanawa 1998, Larson 2003)

dependency on EOS



for γ<1 fragmentation is enhanced → *cluster of low-mass stars*

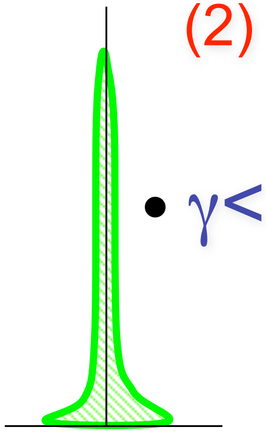
for γ>1 it is suppressed → formation of *isolated massive stars*

(from Li, Klessen, & Mac Low 2003, ApJ, 592, 975)

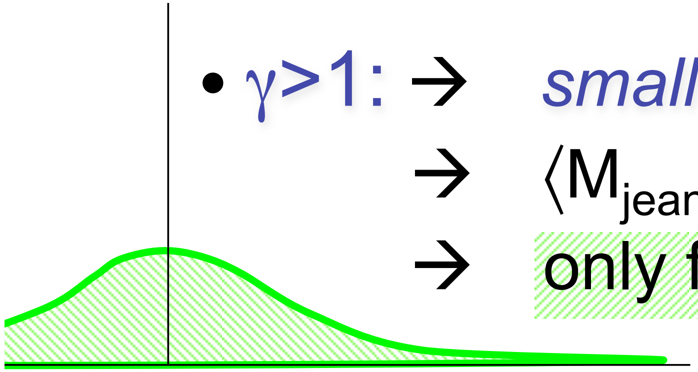
how does that work?

(1) $\mathbf{p} \propto \rho^{\gamma}$ → $\rho \propto \mathbf{p}^{1/\gamma}$

(2) $\mathbf{M}_{\mathbf{jeans}} \propto \gamma^{3/2} \rho^{(3\gamma-4)/2}$



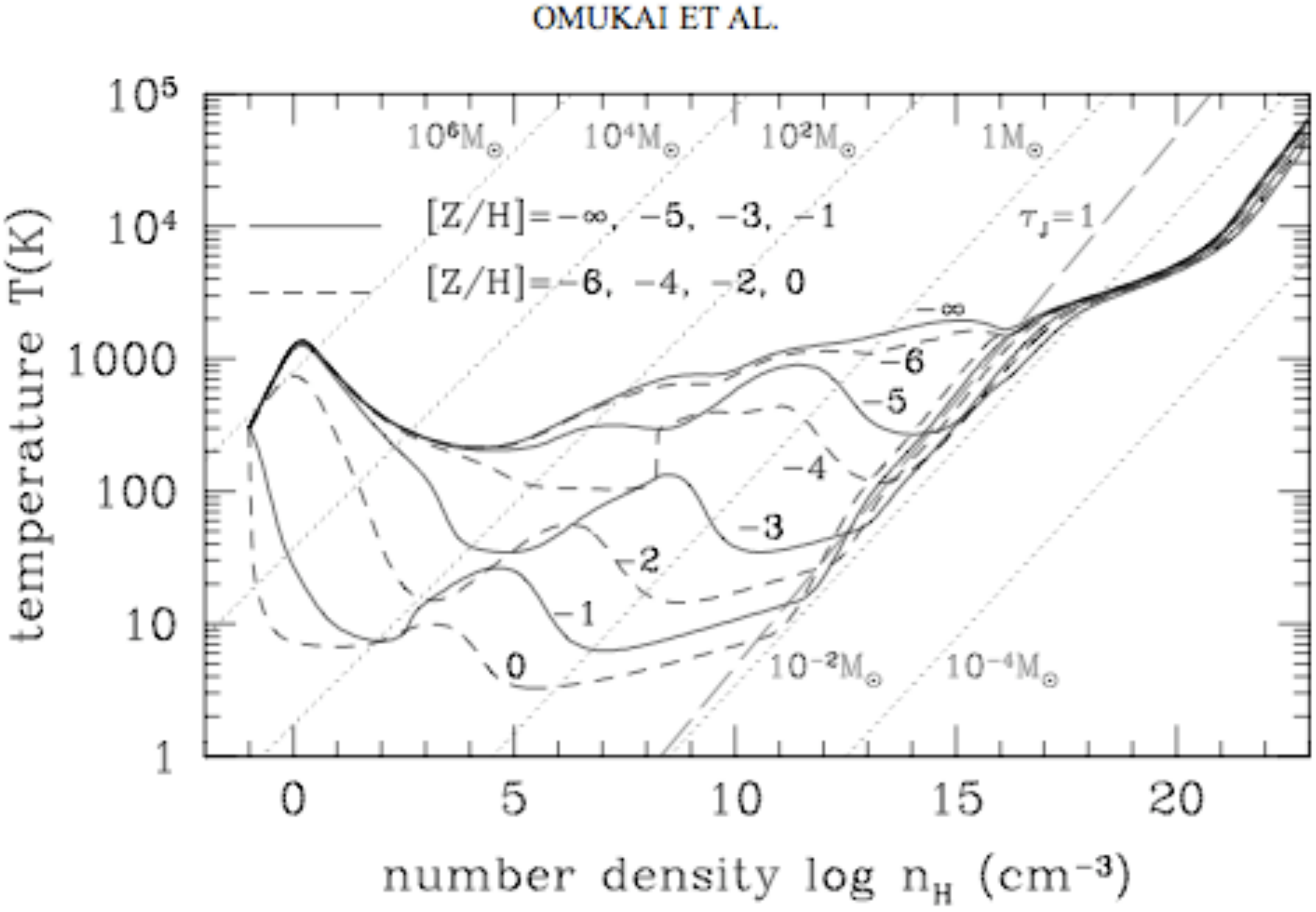
- $\gamma<1$: → *large* density excursion for given pressure
  → $\langle M_{jeans} \rangle$ becomes small
  → number of fluctuations with $M > M_{jeans}$ is large



- $\gamma>1$: → *small* density excursion for given pressure
  → $\langle M_{jeans} \rangle$ is large
  → only few and massive clumps exceed M_{jeans}

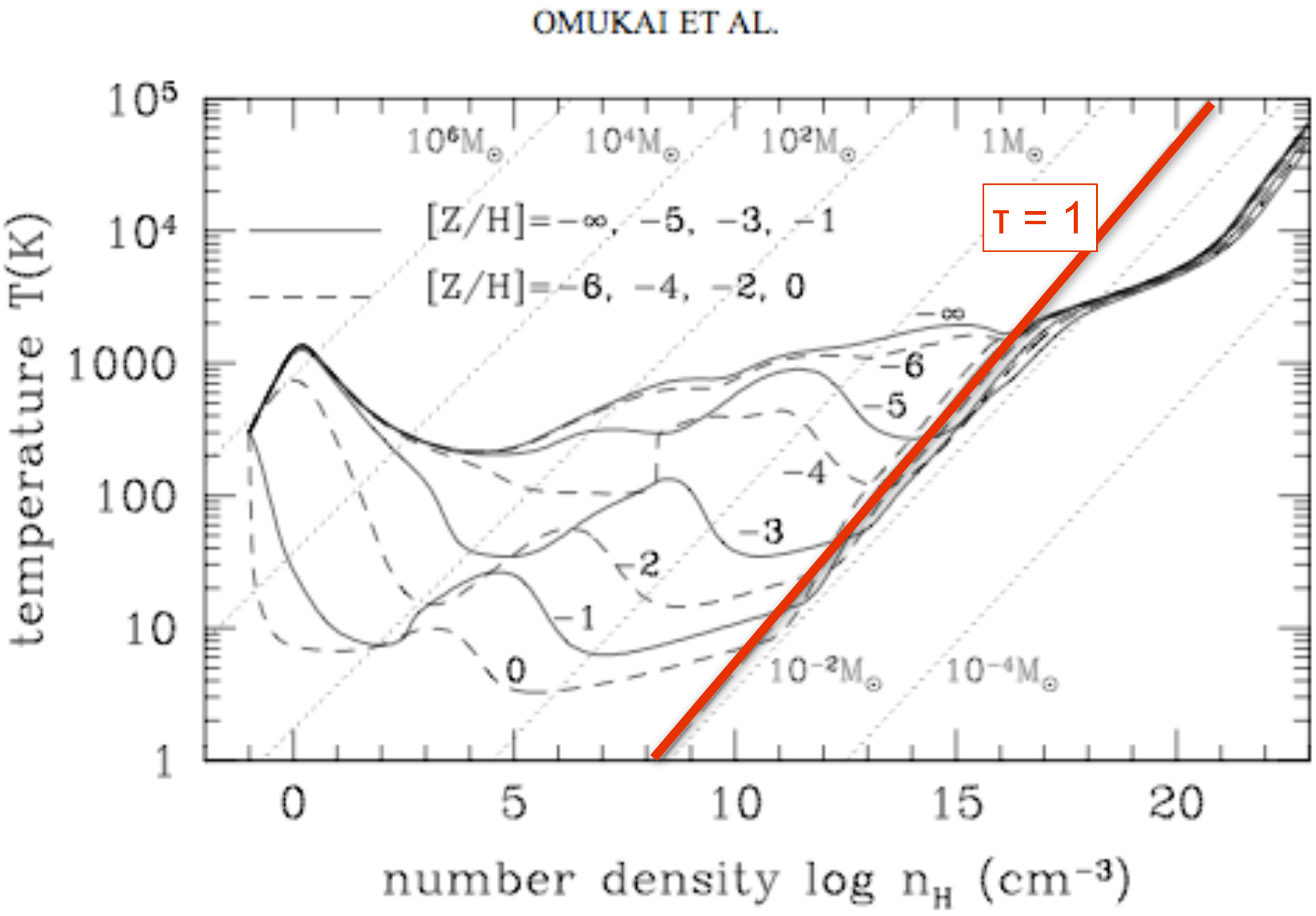
EOS in different environments

EOS as function of metallicity



(Omukai et al. 2005)

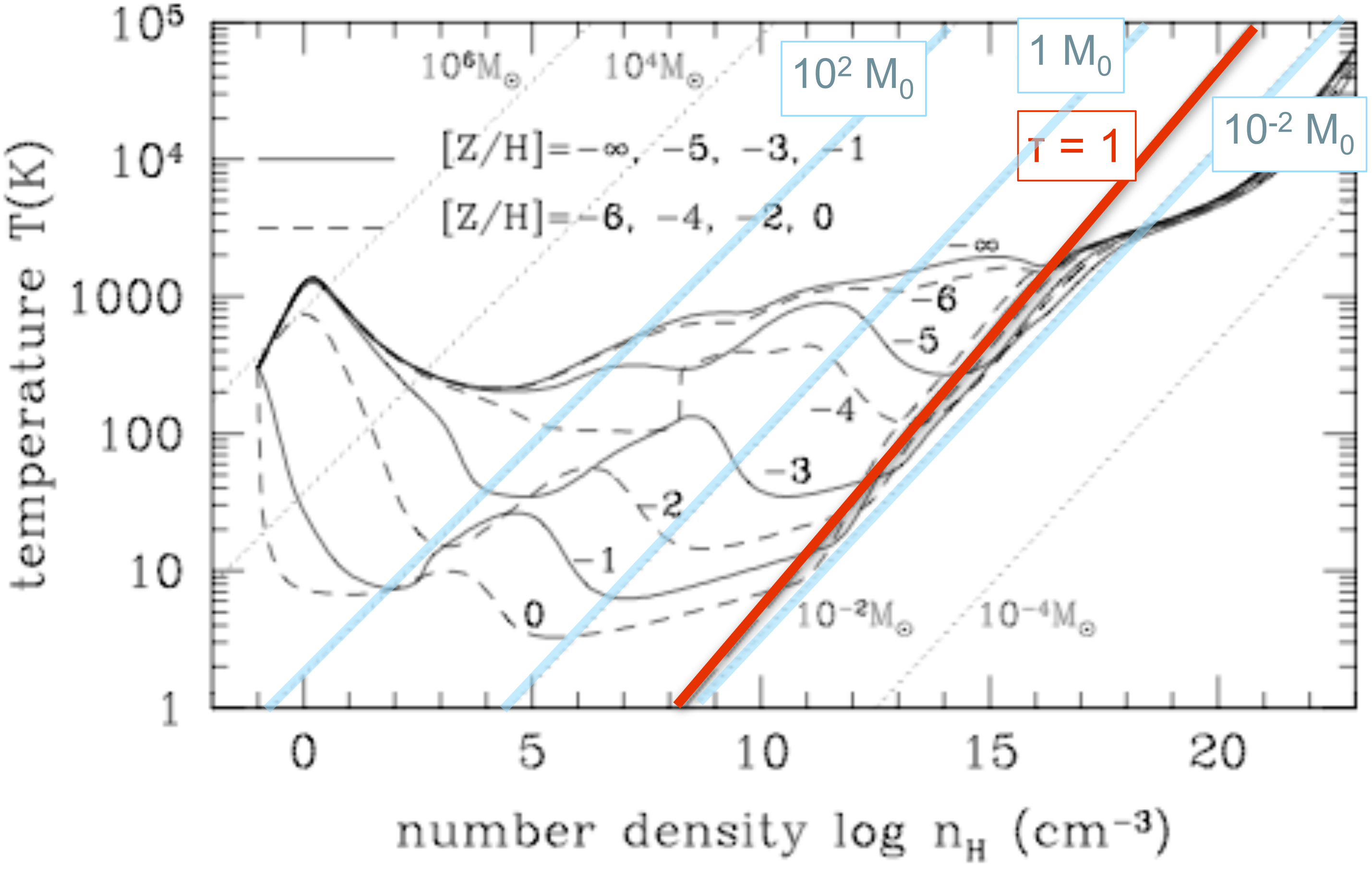
EOS as function of metallicity



(Omukai et al. 2005)

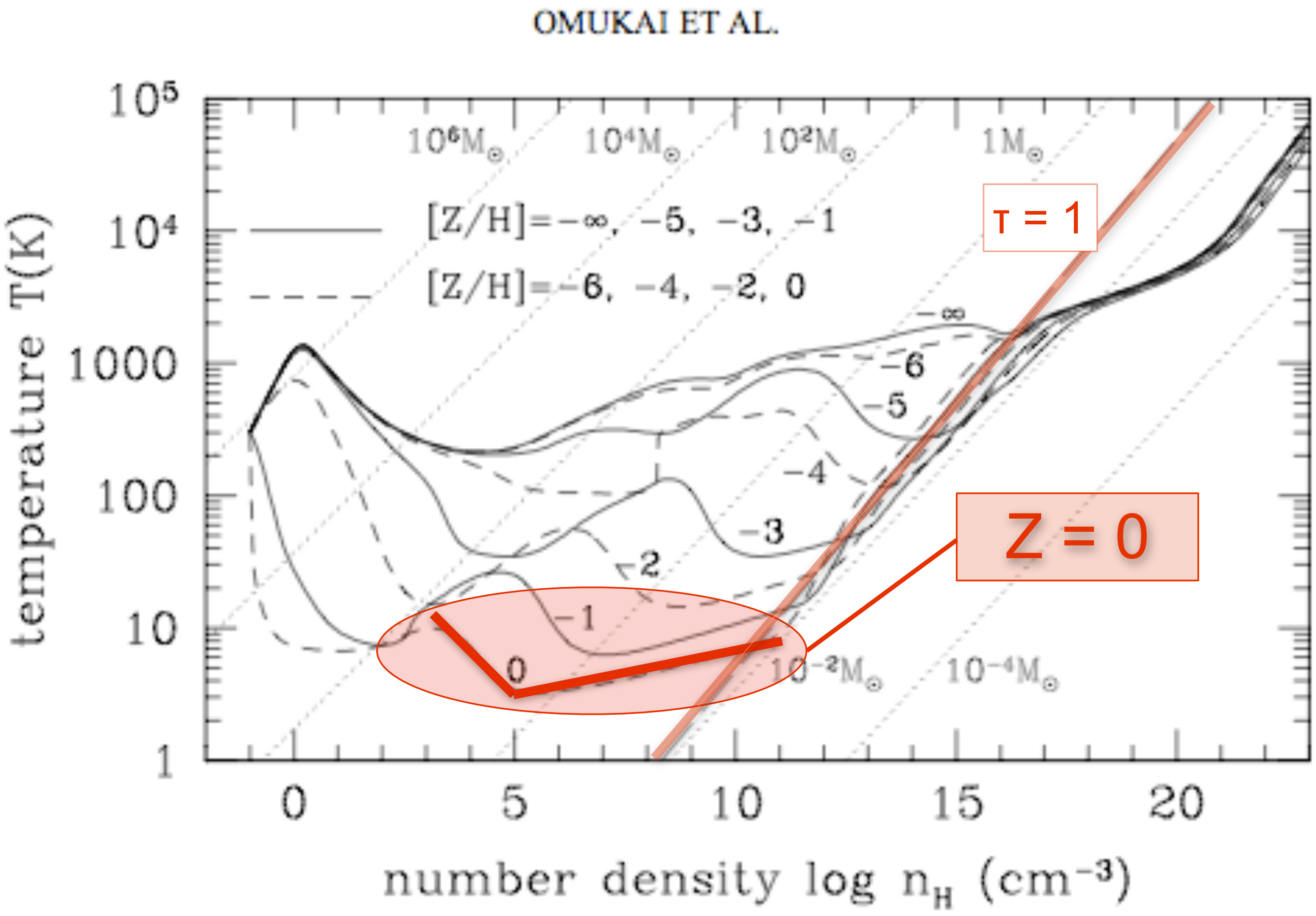
EOS as function of metallicity

OMUKAI ET AL.



(Omukai et al. 2005)

present-day star formation

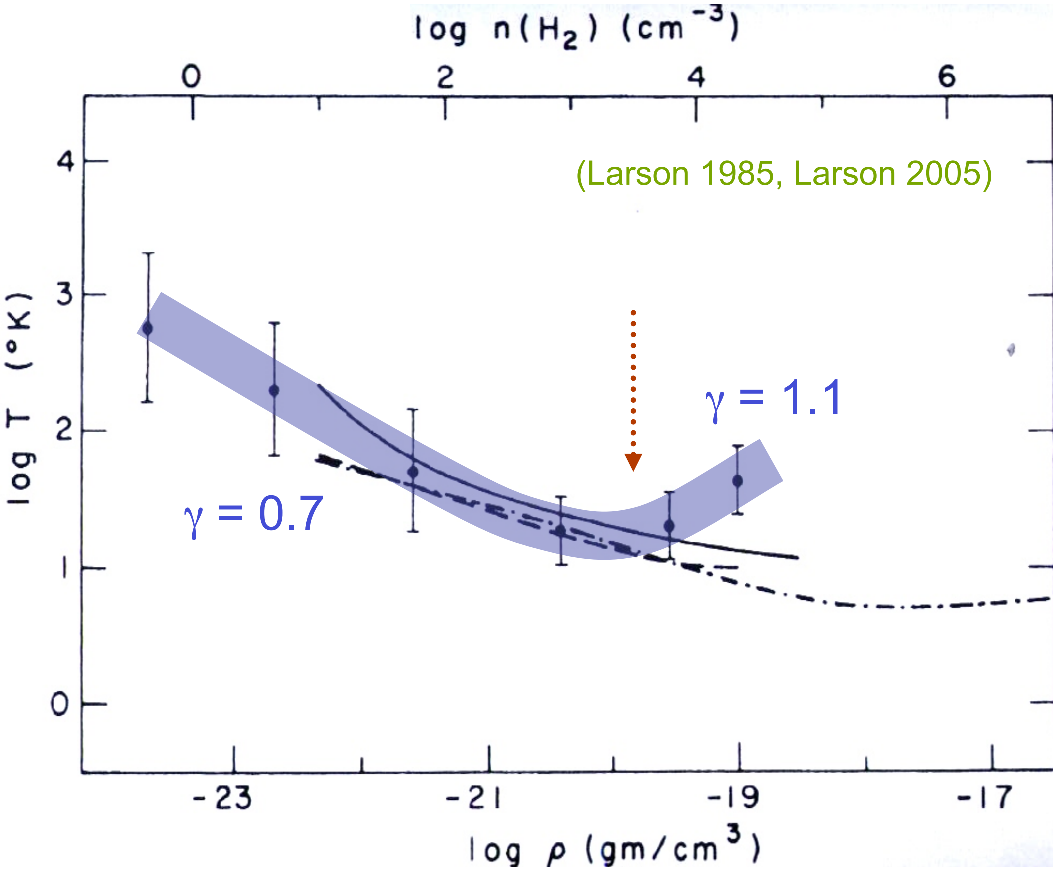


τ = 1

Z = 0

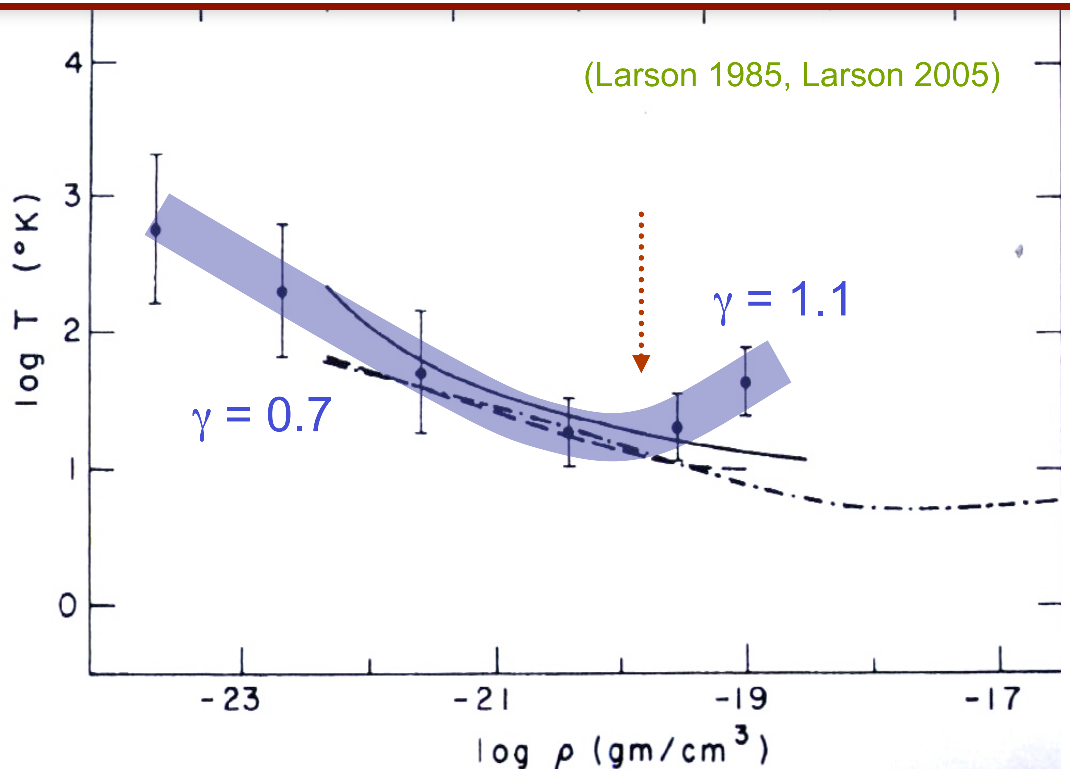
(Omukai et al. 2005, Jappsen et al. 2005, Larson 2005)

present-day star formation



present-day star formation

This kink in EOS is very insensitive to environmental conditions such as ambient radiation field
--> reason for universal for of the IMF? (Elmegreen et al. 2008)

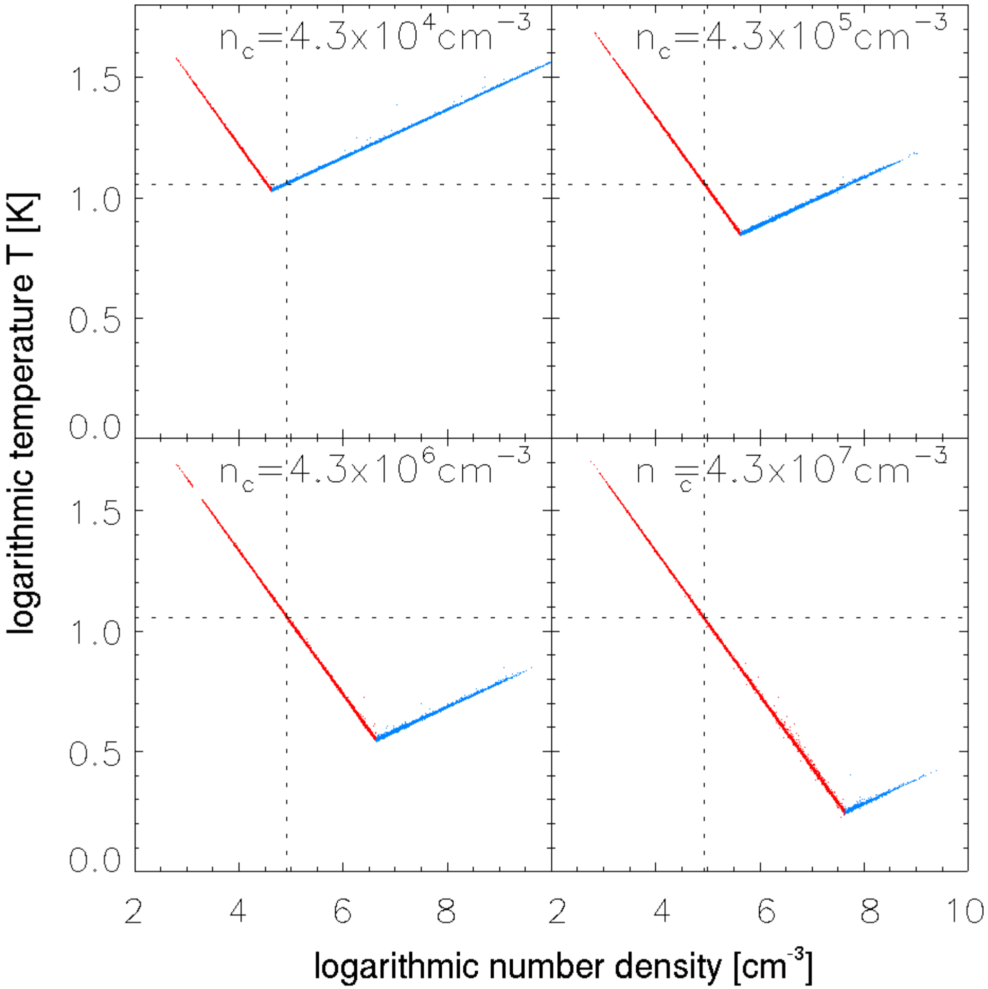


IMF from simple piece-wise polytropic EOS

$\gamma_1 = 0.7$

$\gamma_2 = 1.1$

$$\boxed{T \sim \rho^{\gamma-1}}$$



EOS and Jeans Mass:

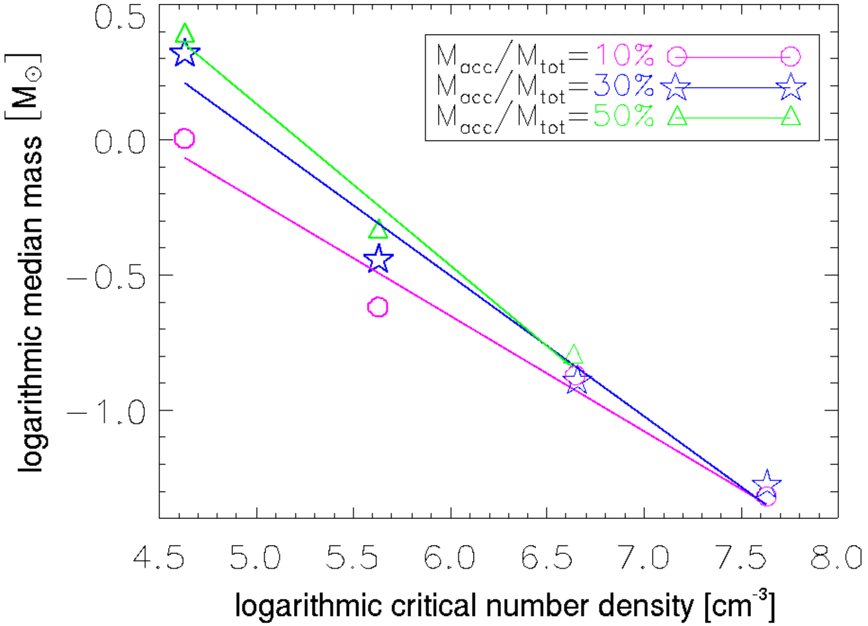
$\mathbf{p} \propto \rho^{\gamma} \rightarrow \rho \propto \mathbf{p}^{1/\gamma}$

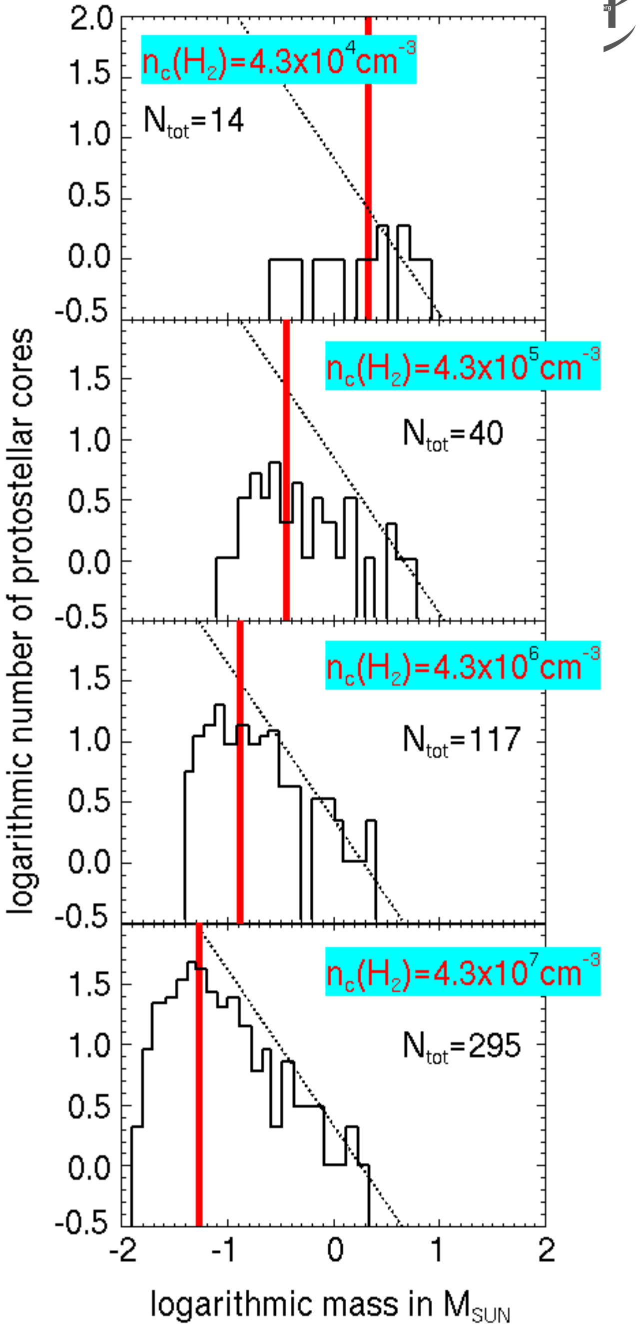
$M_{jeans} \propto \gamma^{3/2} \rho^{(3\gamma-4)/2}$

(Jappsen et al. 2005)

IMF from simple piece-wise EOS

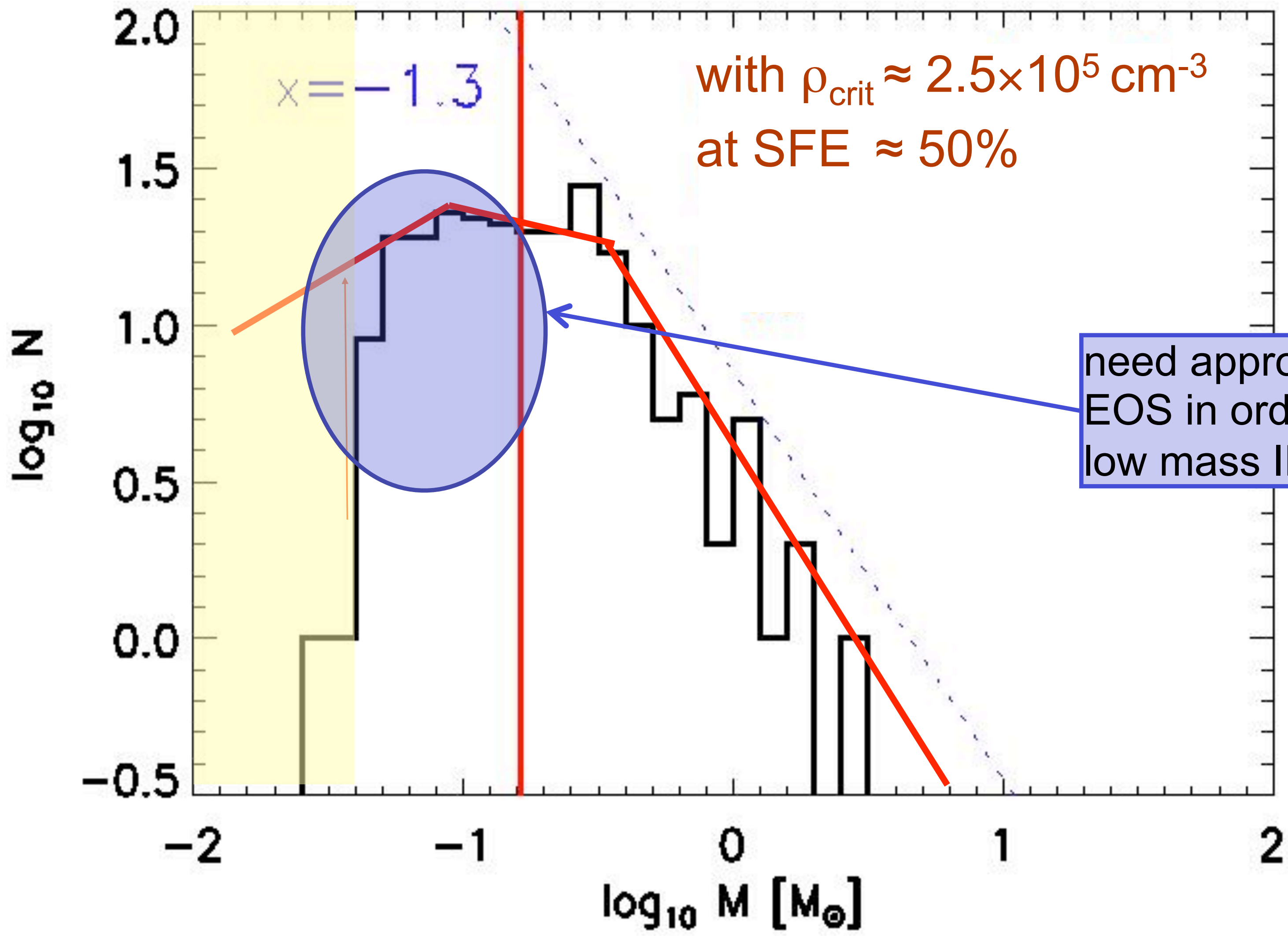
critical density ↑ ⇒ median mass ↓





(Jappsen et al. 2005)

IMF in nearby molecular clouds



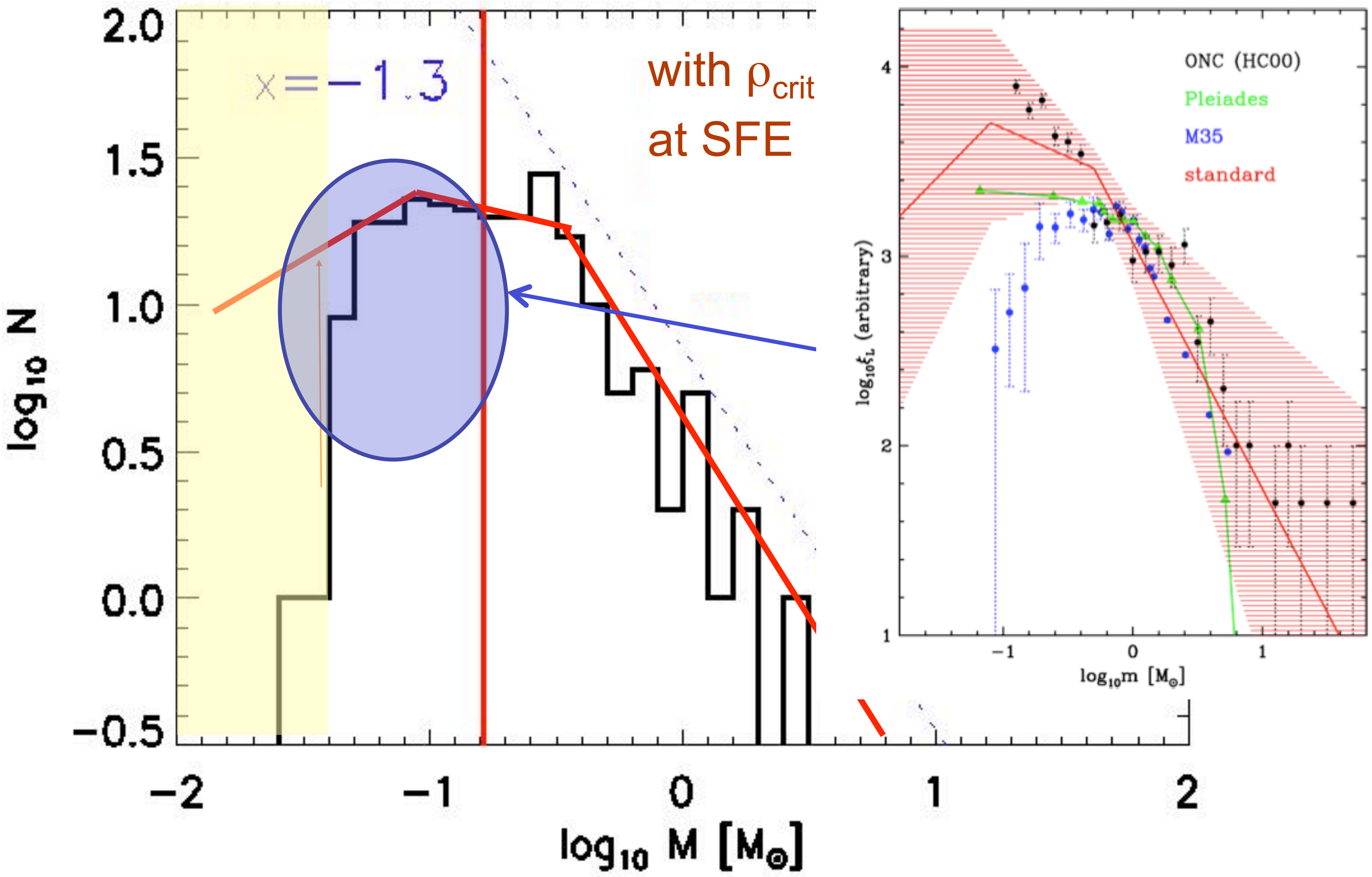
with $\rho_{crit} \approx 2.5\times10^5$ cm^{-3}

at SFE ≈ 50%

need appropriate EOS in order to get low mass IMF right

(Jappsen et al. 2005, A&A, 435, 611)c

IMF in nearby molecular clouds

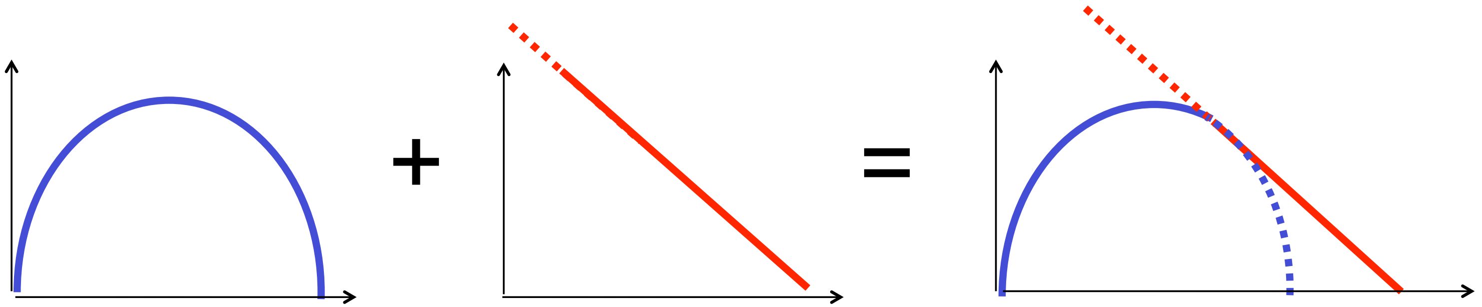


with ρ_{crit}

at SFE

(Jappsen et al. 2005, A&A, 435, 611)c

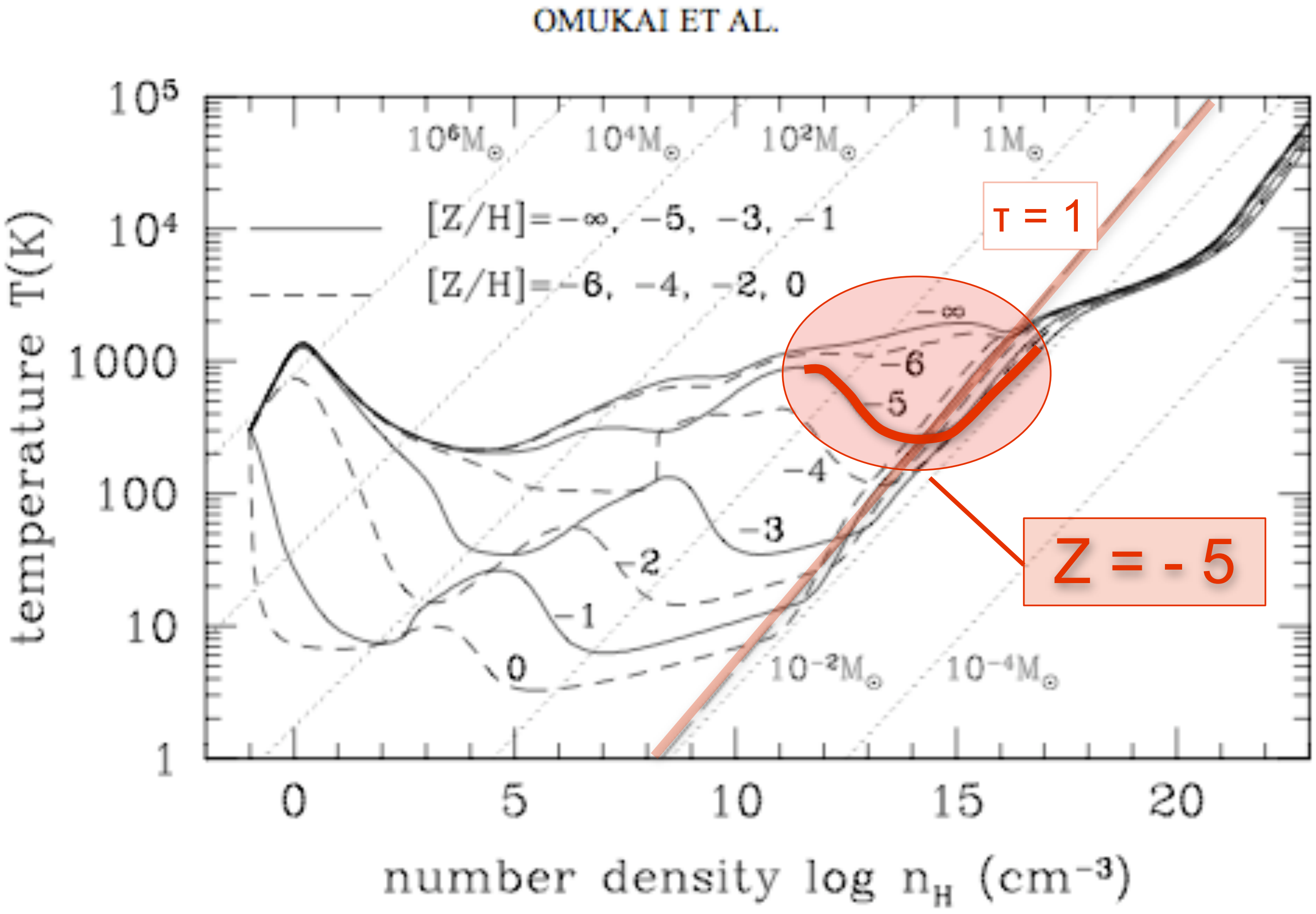
IMF shape and universality



- combine scale free process → *POWER LAW BEHAVIOR*
  - turbulence (Padoan & Nordlund 2002, Hennebelle & Chabrier 2008,2009)
  - gravity in dense clusters (Bonnell & Bate 2006, Klessen 2001)
  - universality: dust-induced EOS kink insensitive to radiation field (Elmegreen et al. 2008)

- with highly stochastic processes → central limit theorem
  → *GAUSSIAN DISTRIBUTION*
  - basically mean thermal Jeans length (or feedback)
  - universality: insensitive to metallicity (Clark et al. 2009)

transition: Pop III to Pop II.5

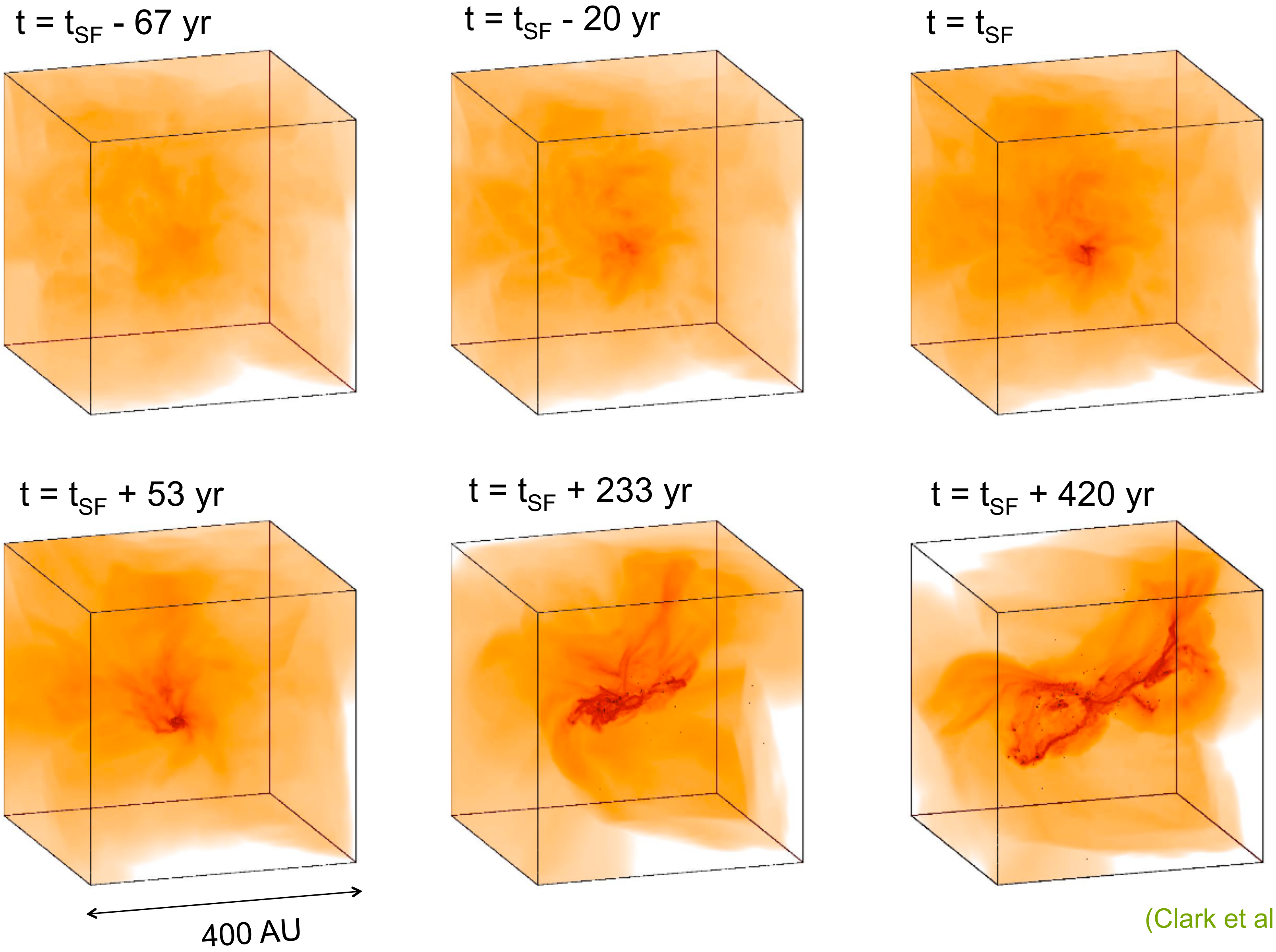


τ = 1

Z = - 5

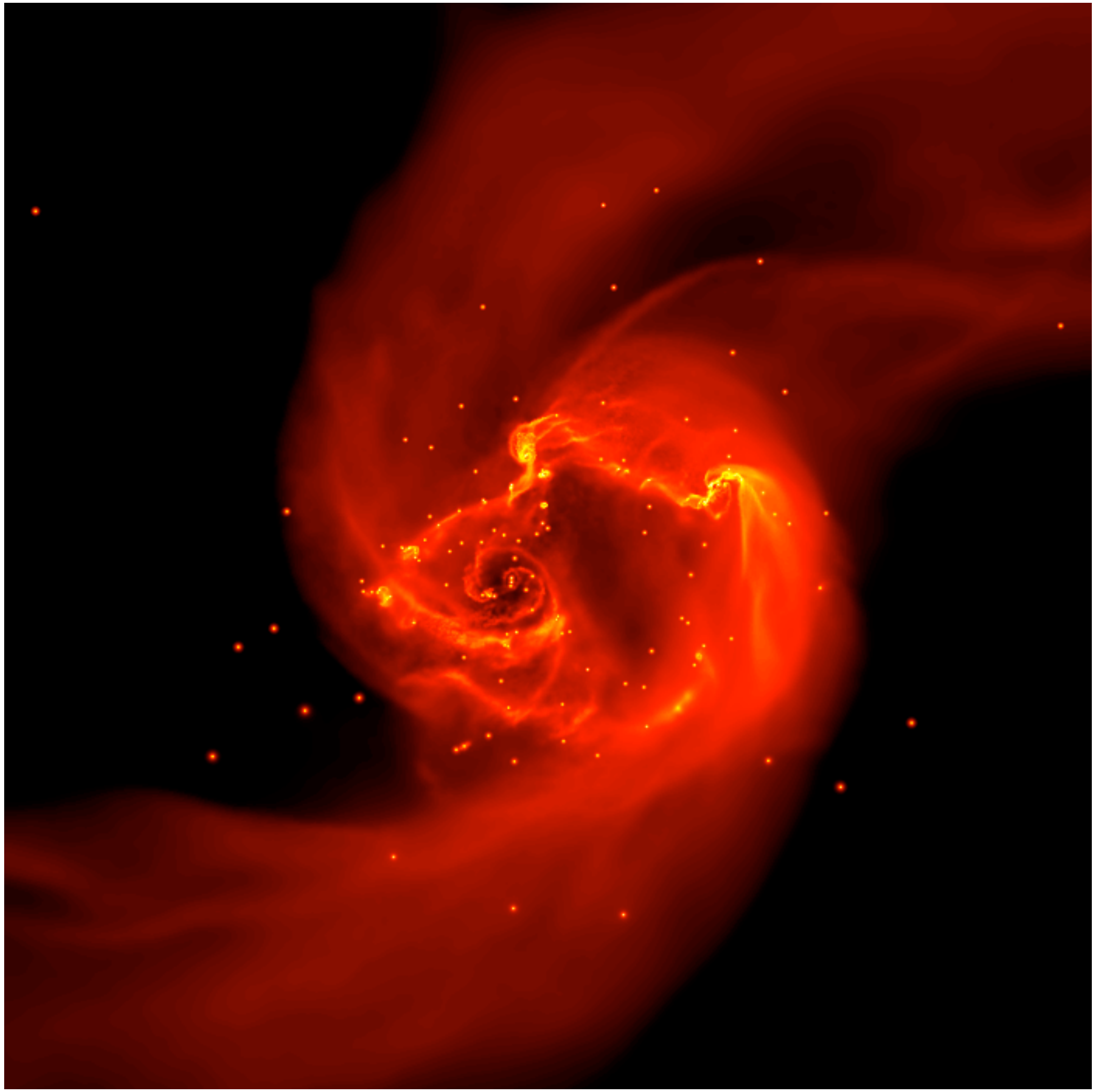
(Omukai et al. 2005)

dust induced fragmentation at $Z=10^{-5}$



(Clark et al. 2007)

dust induced fragmentation at $Z=10^{-5}$

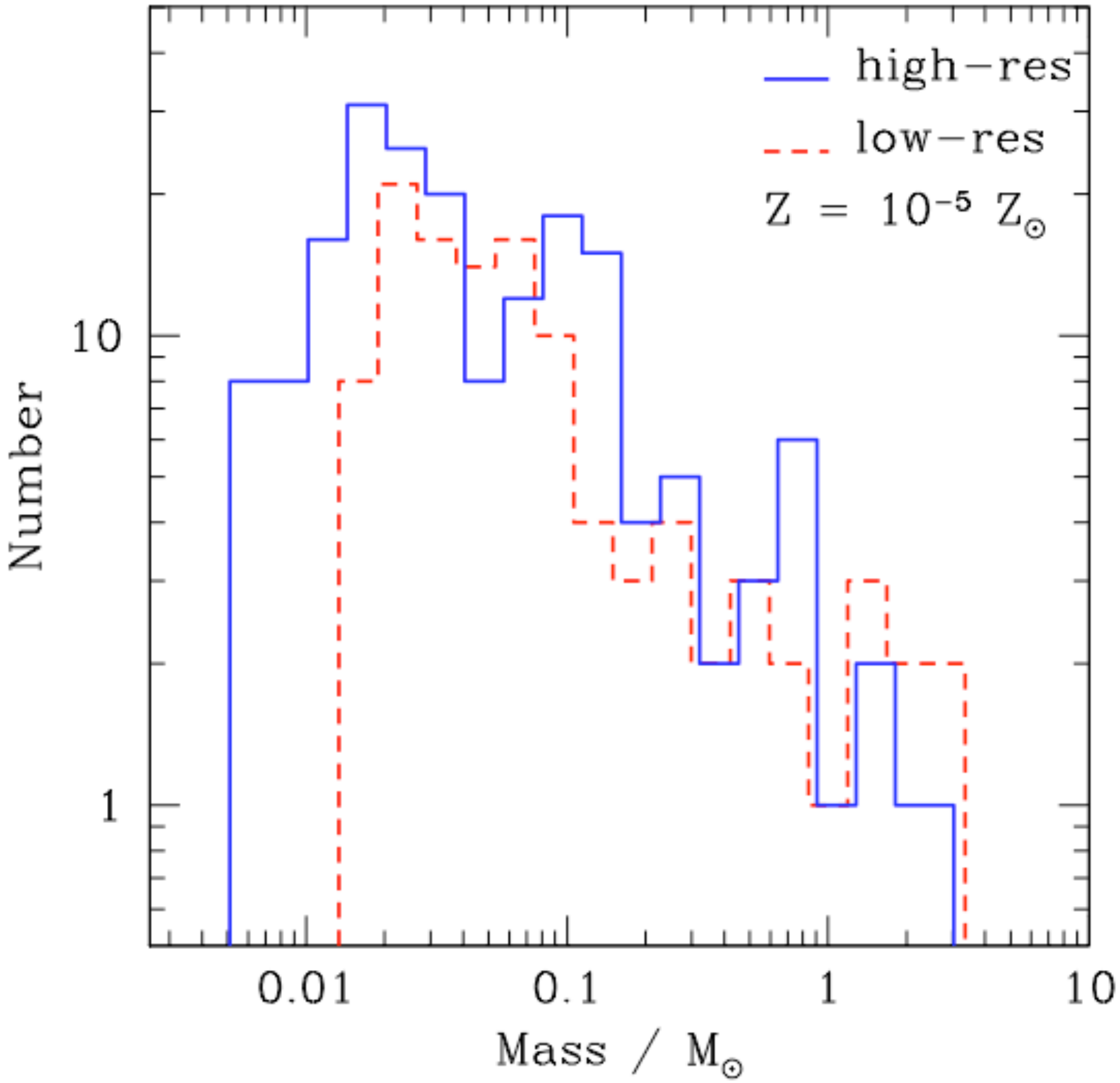


dense cluster of low-mass protostars builds up:

- mass spectrum peaks *below 1* M_{sun}
- cluster VERY dense $n_{stars} = 2.5 \times 10^9\ pc^{-3}$
- fragmentation at density $n_{gas} = 10^{12} - 10^{13}\ cm^{-3}$

(Clark et al. 2008, ApJ 672, 757)

dust induced fragmentation at $Z=10^{-5}$



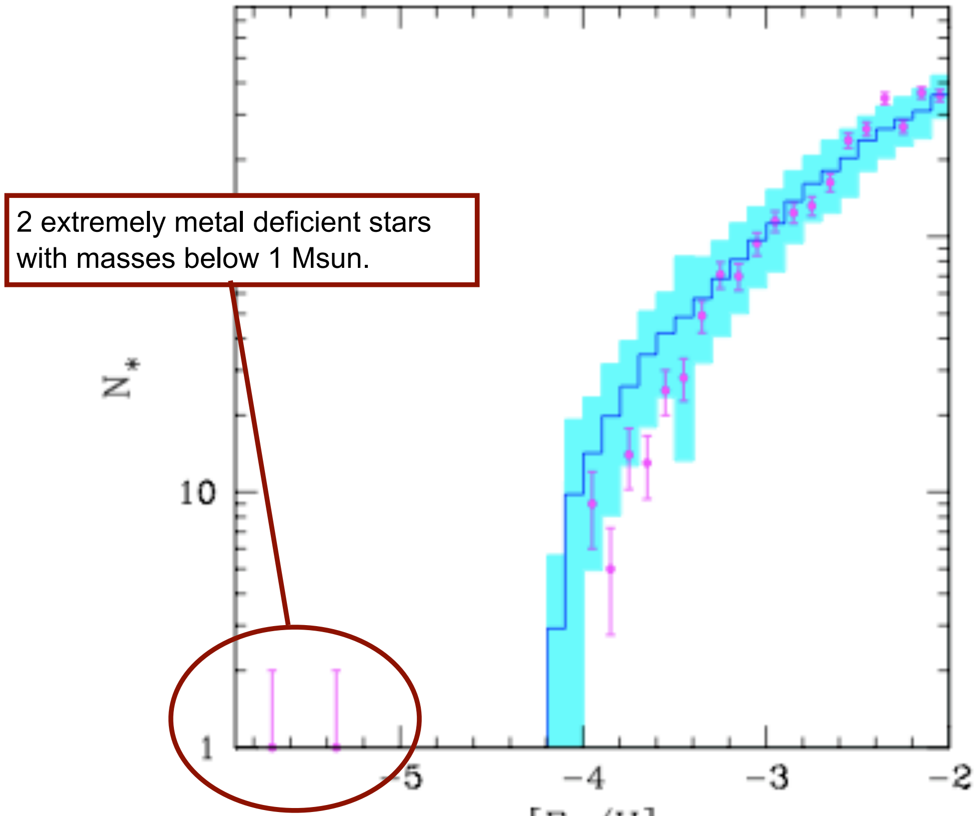
dense cluster of low-mass protostars builds up:

- mass spectrum peaks below 1 M_{sun}
- cluster VERY dense $n_{stars} = 2.5 \times 10^9\ pc^{-3}$

- *predictions:*
  - * low-mass stars with [Fe/H] ~ 10^{-5}
  - * high binary fraction

(Clark et al. 2008)

dust induced fragmentation at $Z=10^{-5}$



(plot from Salvadori et al. 2006, data from Frebel et al. 2005)

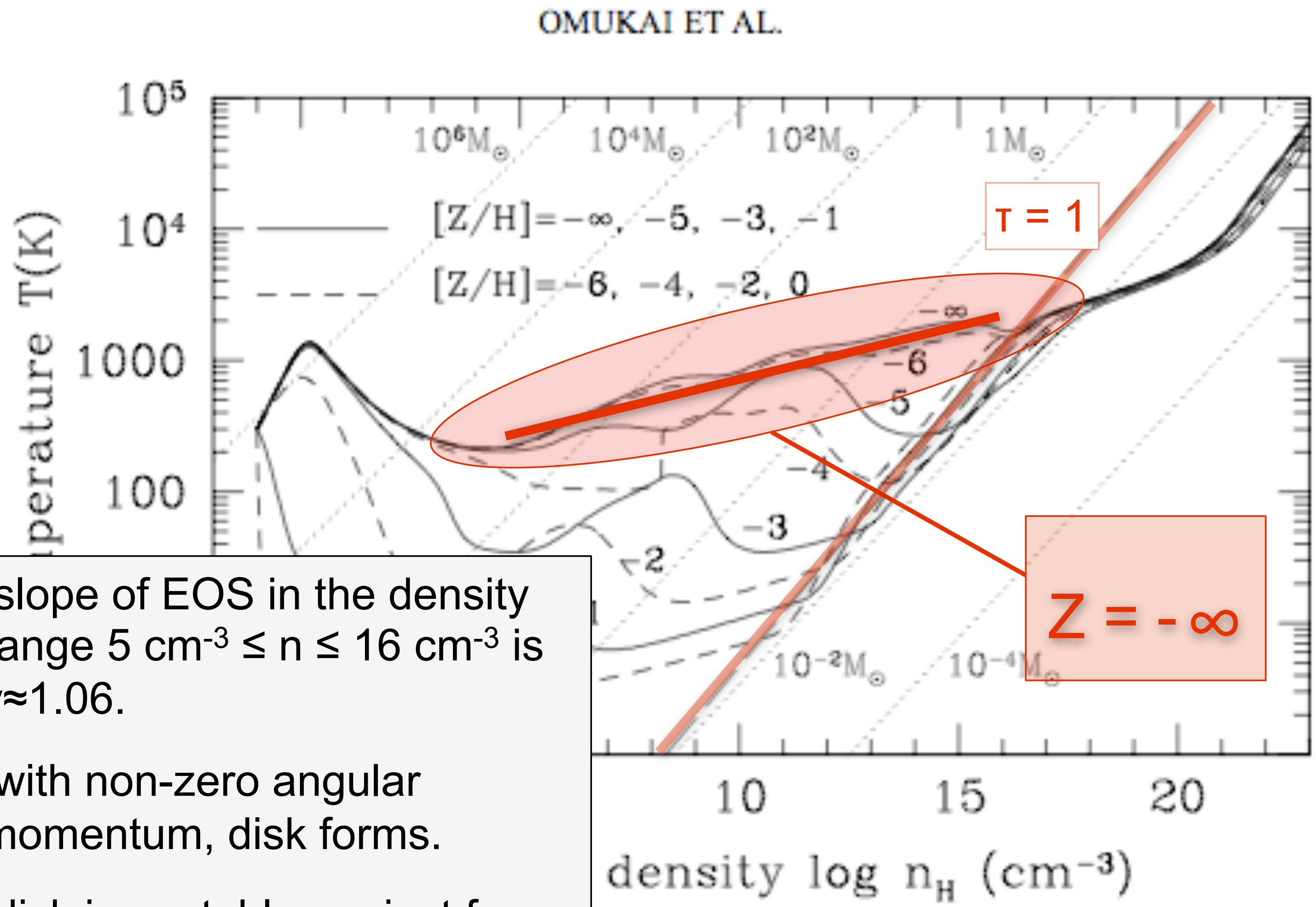
dense cluster of low-mass protostars builds up:

- mass spectrum peaks below 1 M_{sun}
- cluster VERY dense $n_{stars} = 2.5 \times 10^9$ pc^{-3}

- *predictions:*
  - * low-mass stars with [Fe/H] ~ 10^{-5}
  - * high binary fraction

(Clark et al. 2008)

metal-free star formation



τ = 1

Z = - ∞

- slope of EOS in the density range 5 cm^{-3} ≤ n ≤ 16 cm^{-3} is γ≈1.06.
- with non-zero angular momentum, disk forms.
- disk is unstable against fragmentation at high density

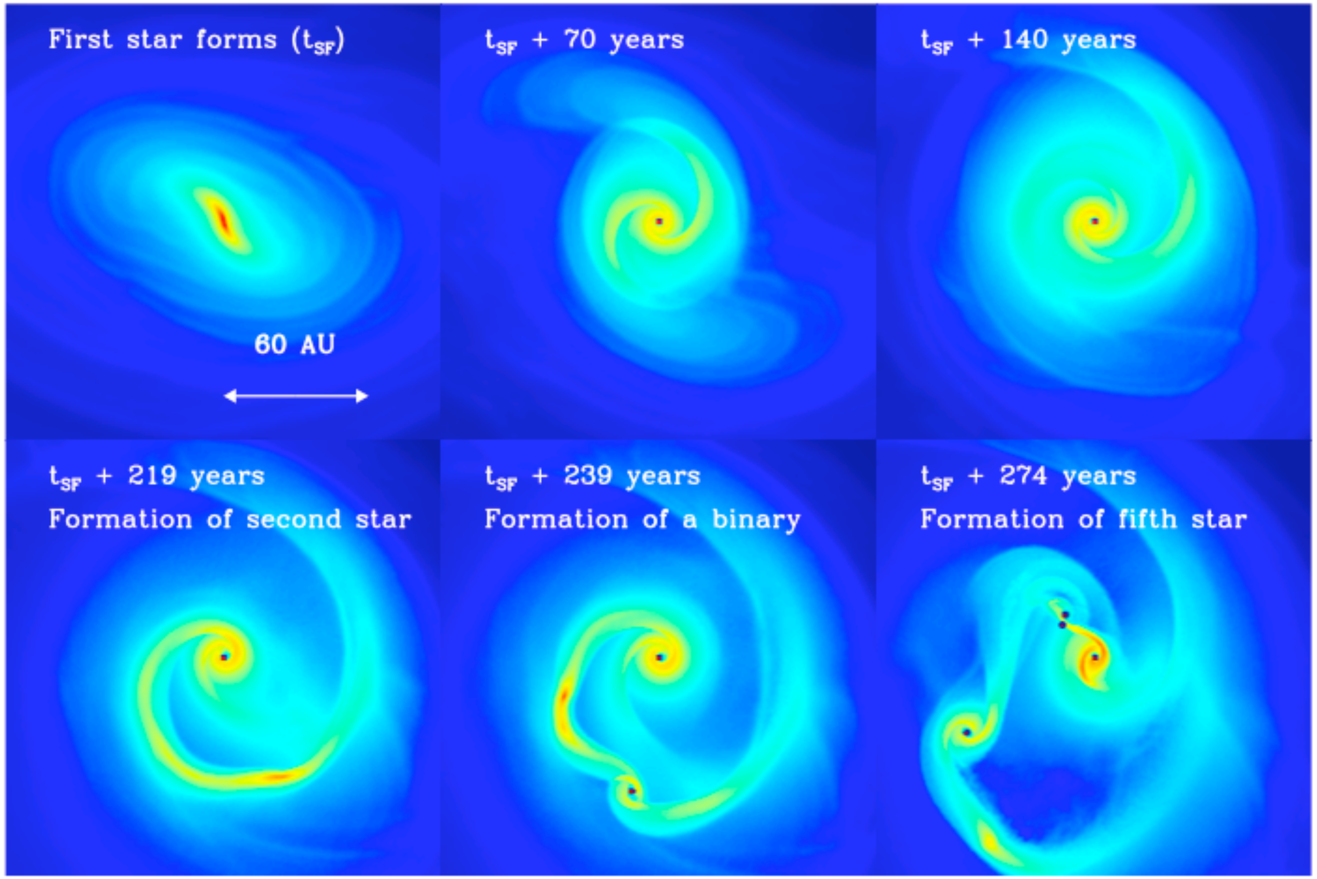


Figure 1: Density evolution in a 180 AU region around the first protostar, showing the build-up of the protostellar disk and its eventual fragmentation. The prominent two-arm spiral structure is caused by the gravitational instability in the disk, and the resulting gravitational torques provide the main source of angular momentum transport that allows disk material to accrete onto the protostar. Eventually, as mass continues to pour onto the disk from the in-falling envelope, the disk becomes so unstable that regions in the spiral arms become self-gravitating in their own right: the disk fragments and a multiple system is formed. The color table is stretched over number densities ranging from 10^{11} (dark blue) to 10^{16} cm^{-3} (red).

(Clark et al. 2010)

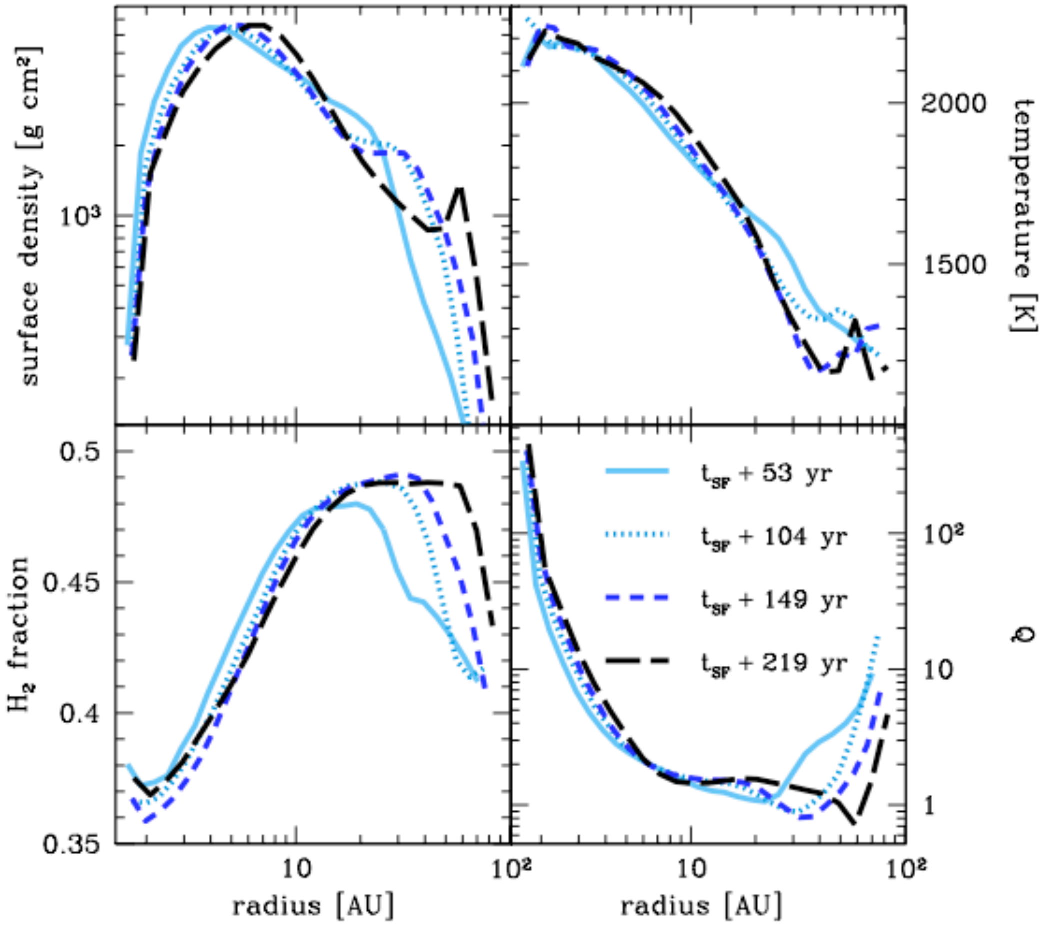


Figure 2: Radial profiles of the disk's physical properties, centered on the first protostellar core to form. The quantities are mass-weighted and taken from a slice through the midplane of the disk. In the lower right-hand plot we show the radial distribution of the disk's Toomre parameter, $Q = c_s\kappa/\pi G\Sigma$, where c_s is the sound speed and κ is the epicyclic frequency. Since our disk is Keplerian, we adopt the standard simplification, and replace κ with the orbital frequency.

(Clark et al. 2010)

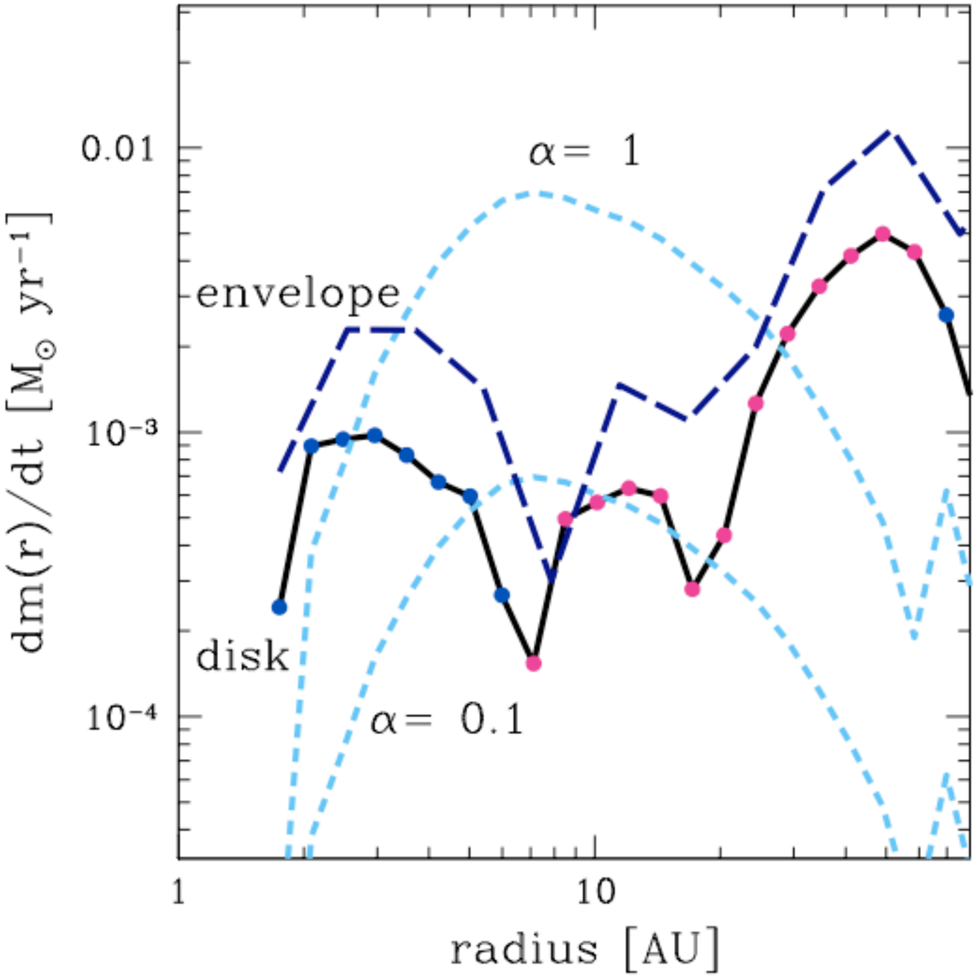
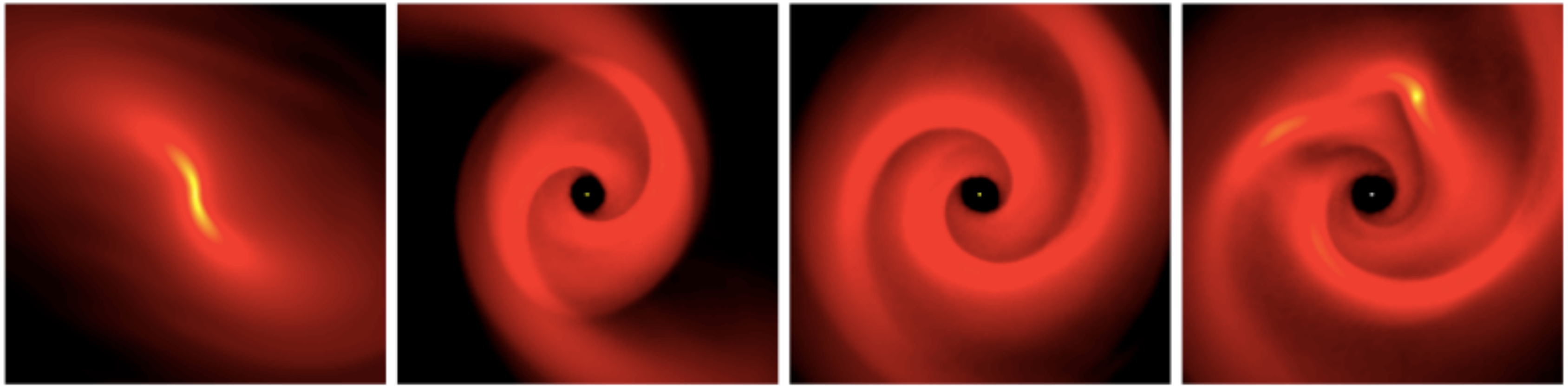


Figure 3: Mass transfer through the disk and in-falling envelope as a function of radius from the central protostar at the onset of disk fragmentation. In the case of the disk we have denoted annuli that are moving towards the protostar with blue dots, and those moving away in pink. The light blue dashed lines show the accretion rates expected from an 'alpha' (thin) disk model, where $\dot{M}(r) = 3\,\pi\,\alpha\,c_{\rm s}(r)\,\Sigma(r)\,H(r)$, with two global values of alpha and where $c_{\rm s}(r)$, $\Sigma(r)$, and $H(r)$ are (respectively) the sound speed, surface density and disk thickness at radius r.

(Clark et al. 2010)

primordial star formation



- first star formation is not less complex than present-day star formation
- brave claim: *all* Pop III stars form in multiple systems
- even braver claim: some Pop III stars fall in the mass range < 0.5 $M_\odot$ ---> they should still *be around*!!!!

summary

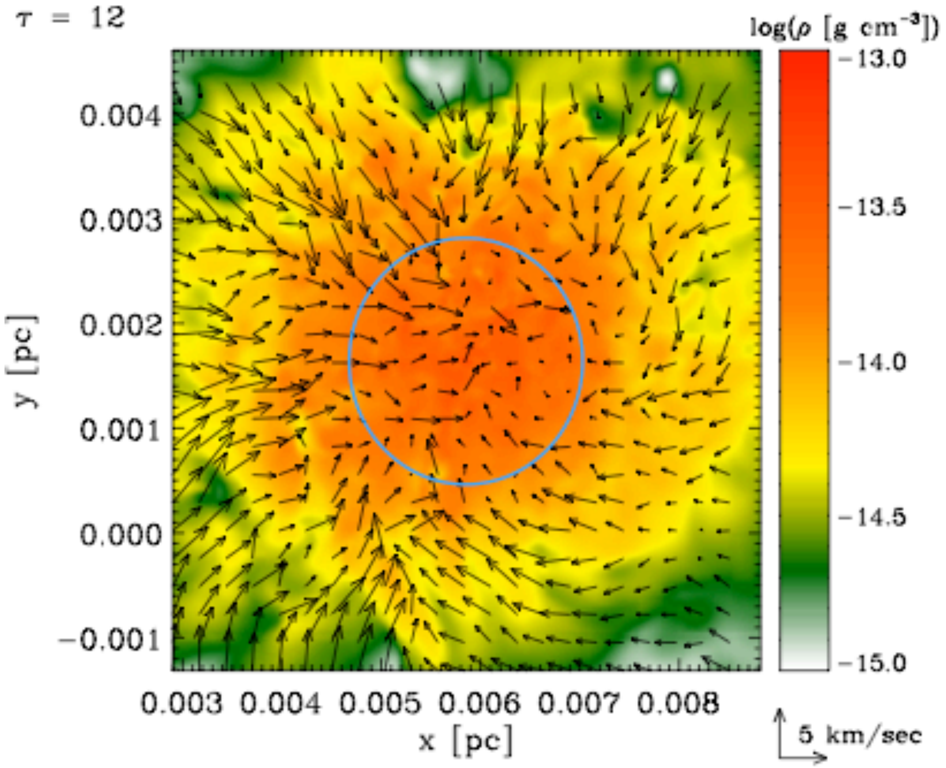
Beispiel

how to make strong B-fields

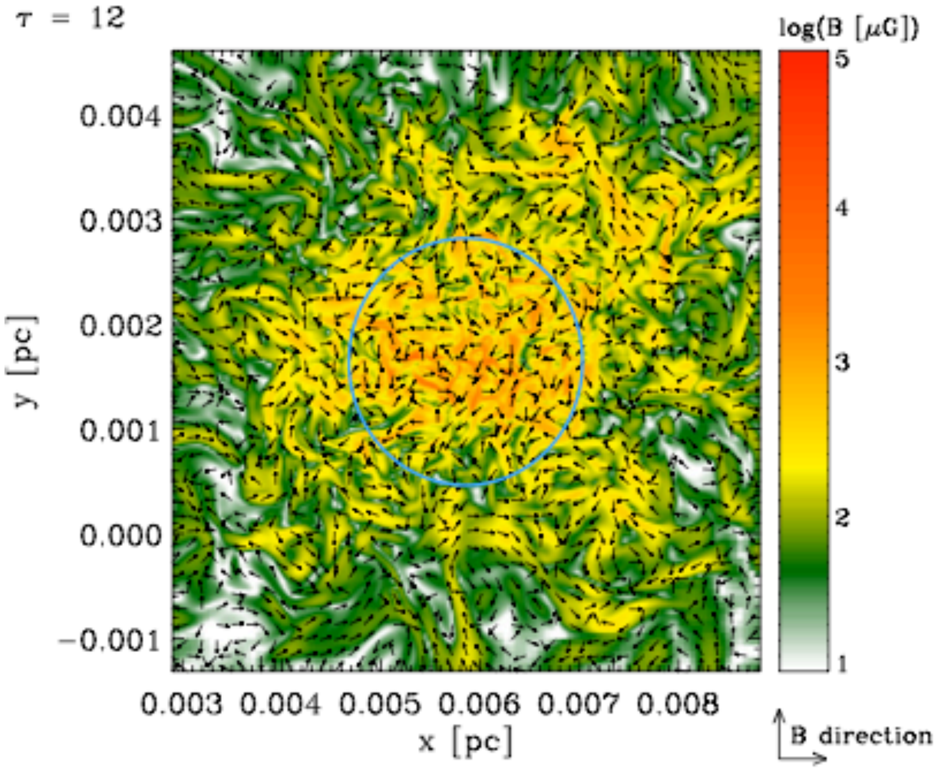
- we know the universe is magnetized (now)
- knowledge about B-fields in the high-redshift universe is extremely uncertain
  - inflation / QCD phase transition / Biermann battery / Weibel instability
- they are thought to be extremely small
- however, THIS MAY BE WRONG!

small-scale turbulent dynamo

- *idea:* the small-scale turbulent dynamo can generate strong magnetic fields from very small seed fields
- *approach:* model collapse of primordial gas ---> formation of the first stars in low-mass halo at redshift z ~ 20
- *method:* solve ideal MHD equations with very high resolution
  - grid-based AMR code FLASH
  - resolution up to 128^3 cells per Jeans volume (effective resolution 65536^3 cells)

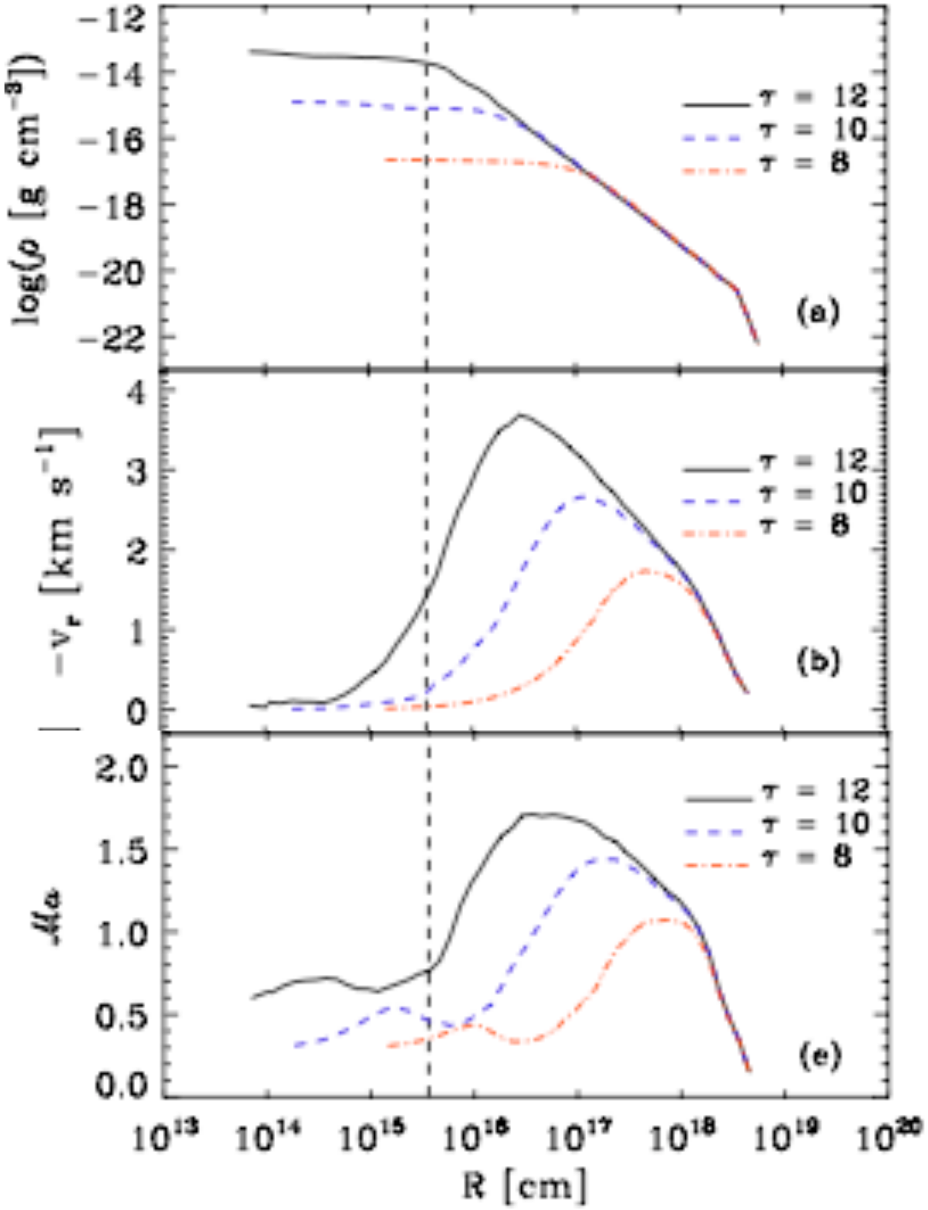


magnetic field structure



density structure

(Sur et al. 2010, ApJ, 721, L734)

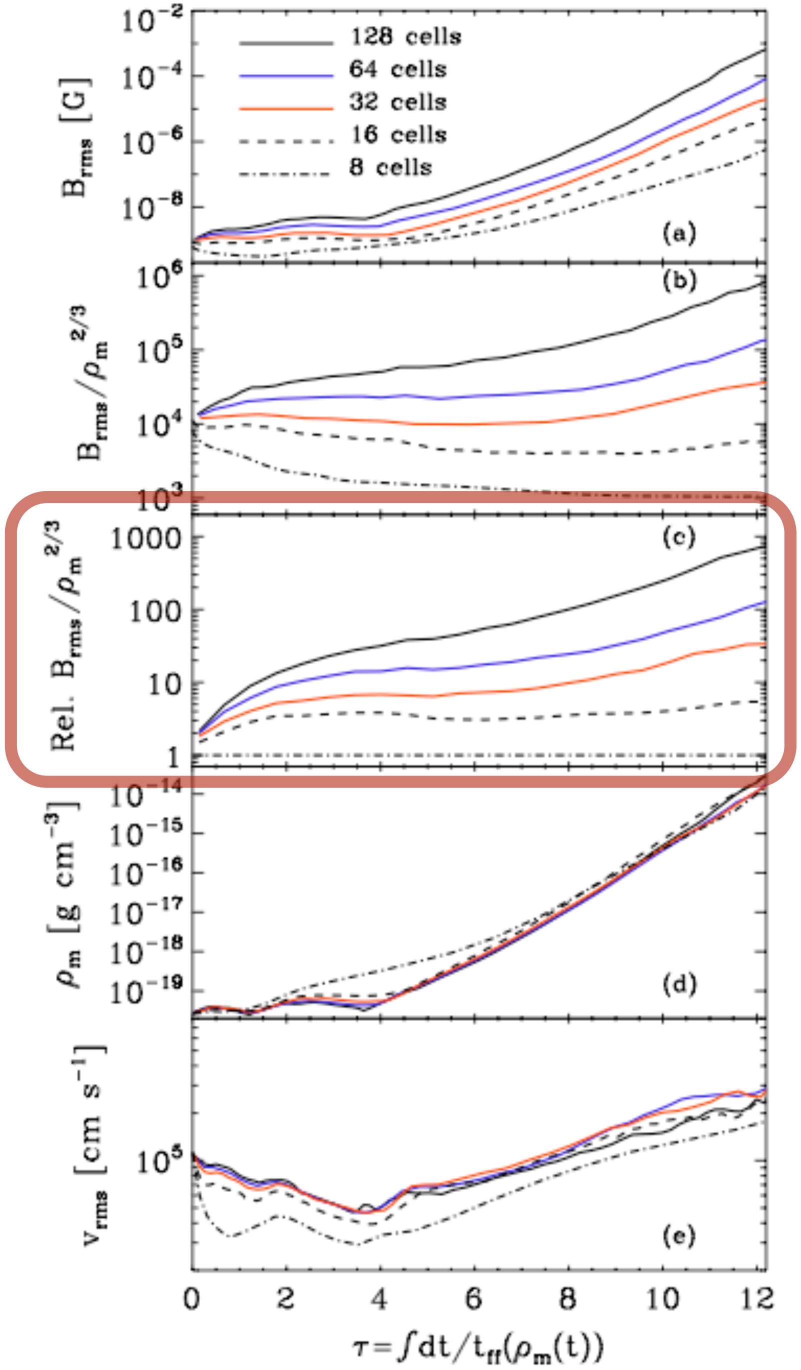


radial density profile

radial velocity profile

Mach number profile

(Sur et al. 2010, ApJ, 721, L734)



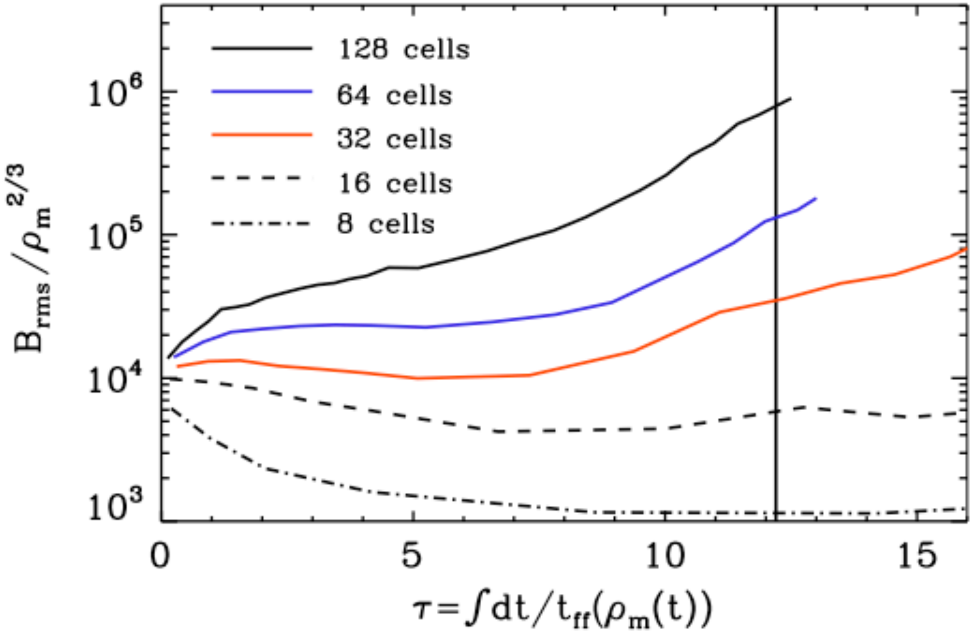


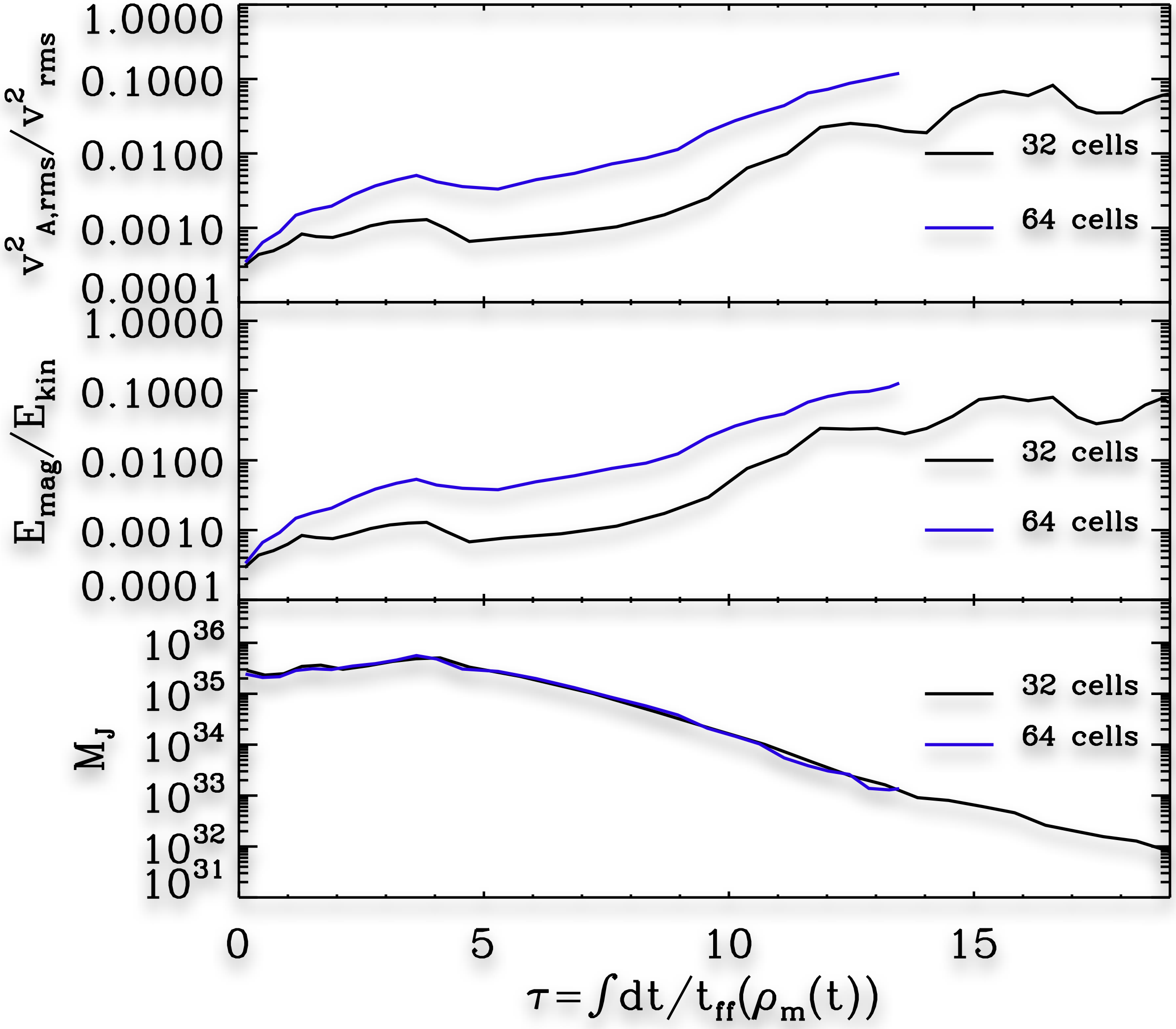
Figure 4. Illustration of the minimum resolution criterion required to capture the growth of the magnetic field due to small-scale dynamo action. The dynamo begins to be observed for simulations where λ_J is resolved by a minimum of 32 cells. Simulations performed with the Jeans length resolved by either 8 cells or 16 cells are decaying in nature with weak fluctuations. The vertical line indicates the values of $B_{\rm rms}/\rho_{\rm m}^{2/3}$ obtained in the different resolution runs at $\tau = 12.2$.

Field amplification during first collapse seems unavoidable.

QUESTION: What is the saturation value? Can the field reach dynamically important strength?

(Sur et al. 2010, ApJ, 721, L734)

first attempts to calculate the saturation level.



Stellare Massenfunktion

- Verteilung der Sternmassen hängt ab von
  - turbulenten Anfangsbedingungen,
    --> Massenspektrum prästellarer Kerne
  - Kollaps und Wechselwirkung prästellarer Kerne,
    --> kompetitive Akkretion & N-Körperdynamik
  - thermodynamischen Eigenschaften des Gases.
    --> Gleichgewicht von Heiz- und Kühlprozessen
    --> Zustandsgleichung (bestimmt welche Wolkenkerne kollabieren)
  - (Proto) stellares Feedback beendet Sternbildung.
    ionisierende Strahlung, bipolare Ausflüsse, Sternwinde, SN

(e.g. Larson 2003, Prog. Rep. Phys.; Mac Low & Klessen, 2004, Rev. Mod. Phys, 76, 125 - 194)

Modell der Orion Wolke

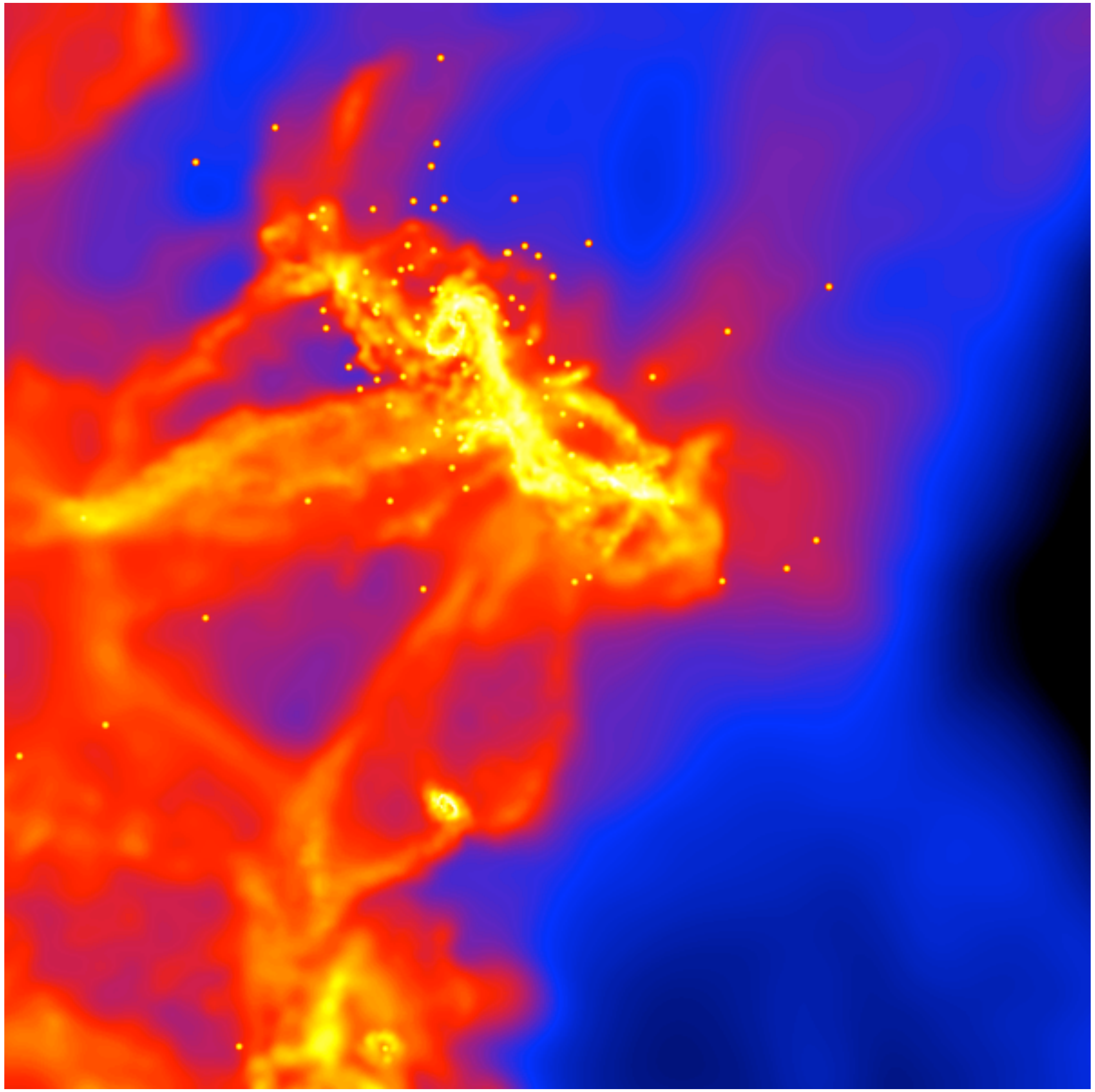
„model“ of Orion cloud: 15.000.000 SPH particles, 10^4 M_{sun} in 10 pc, mass resolution 0,02 M_{sun}, forms ~2.500 „stars“ (sink particles)

isothermal EOS, top bound, bottom unbound

has clustered as well as distributed „star“ formation

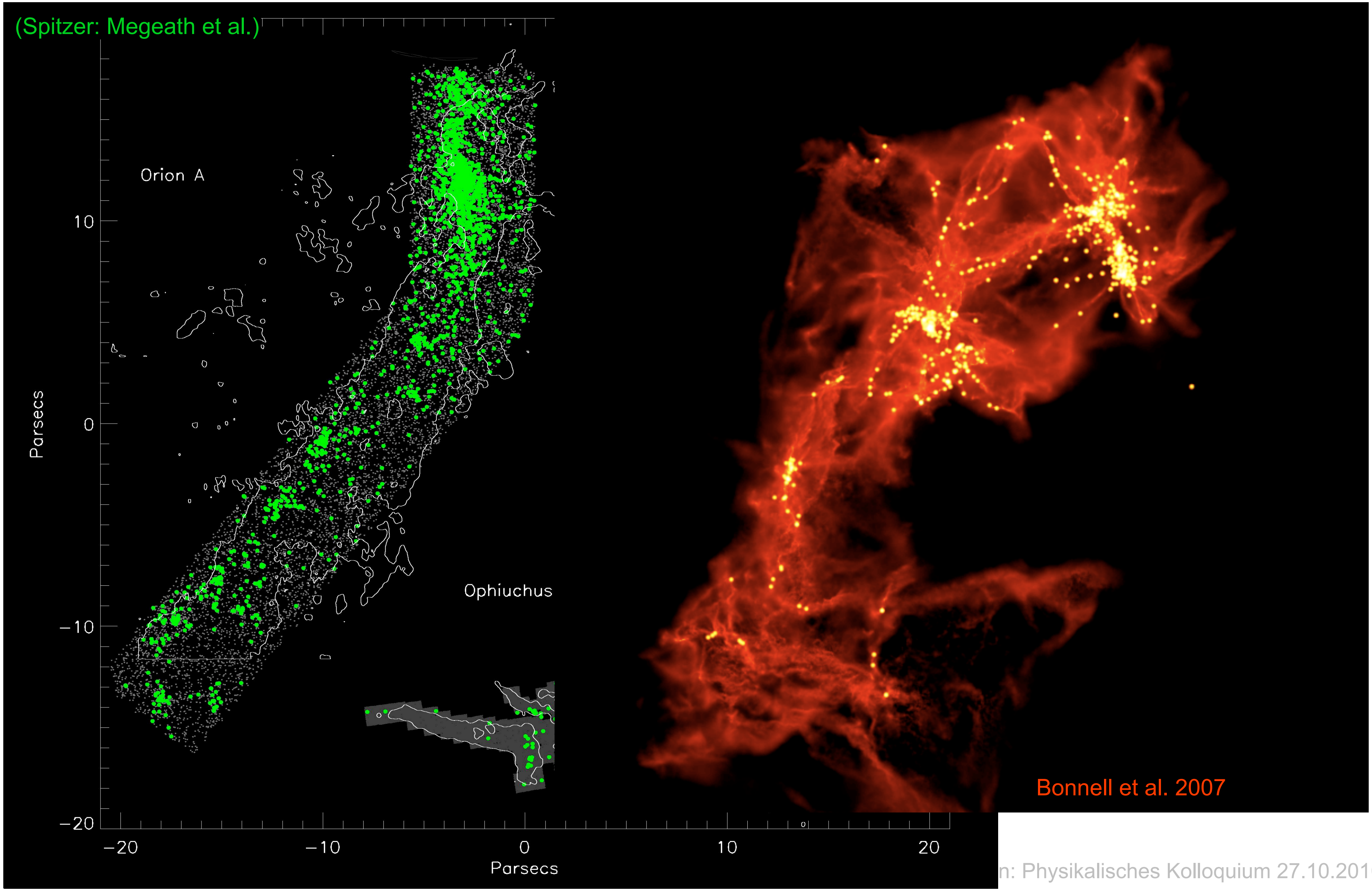
efficiency varies from 1% to 20%

develops full IMF
(distribution of sink particle masses)



(Bonnell et al. 2007)

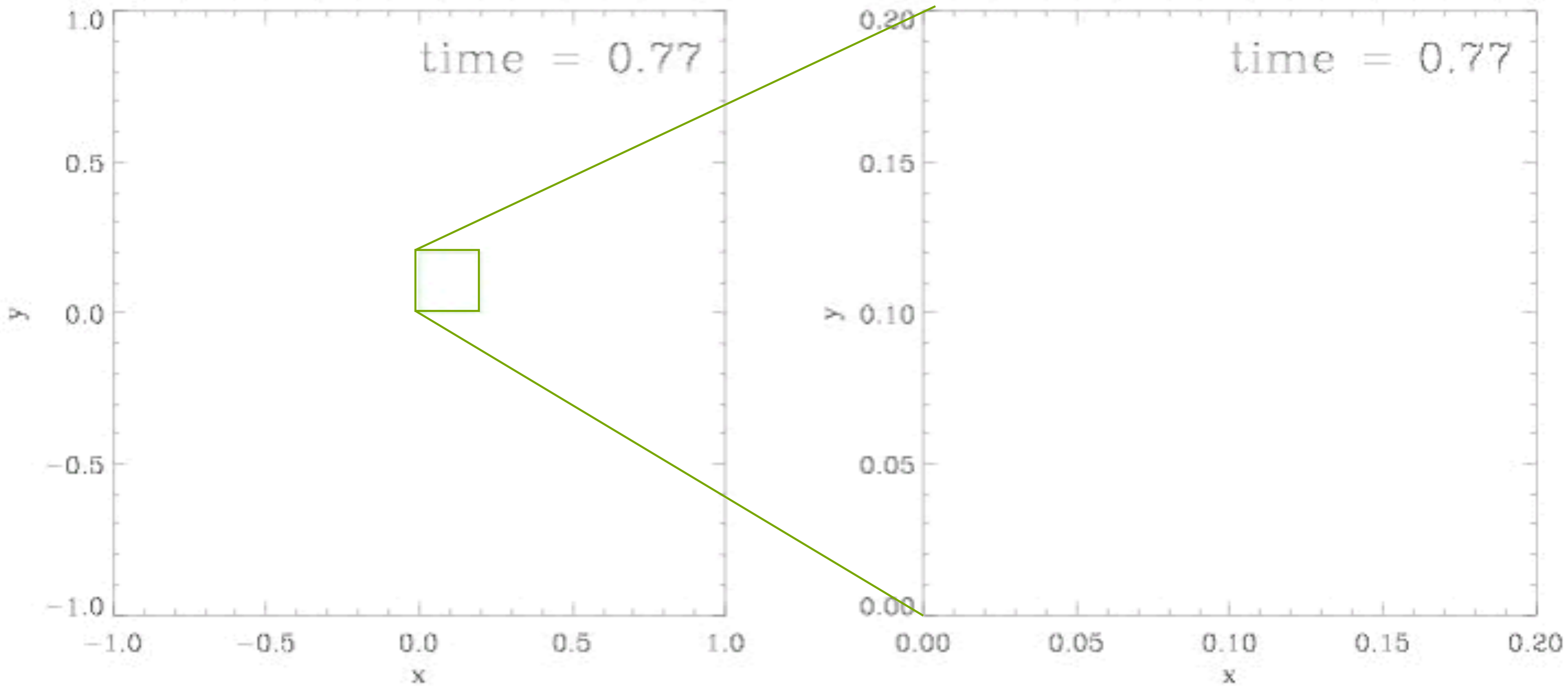
Modell der Orion Wolke


(Spitzer: Megeath et al.)

Bonnell et al. 2007

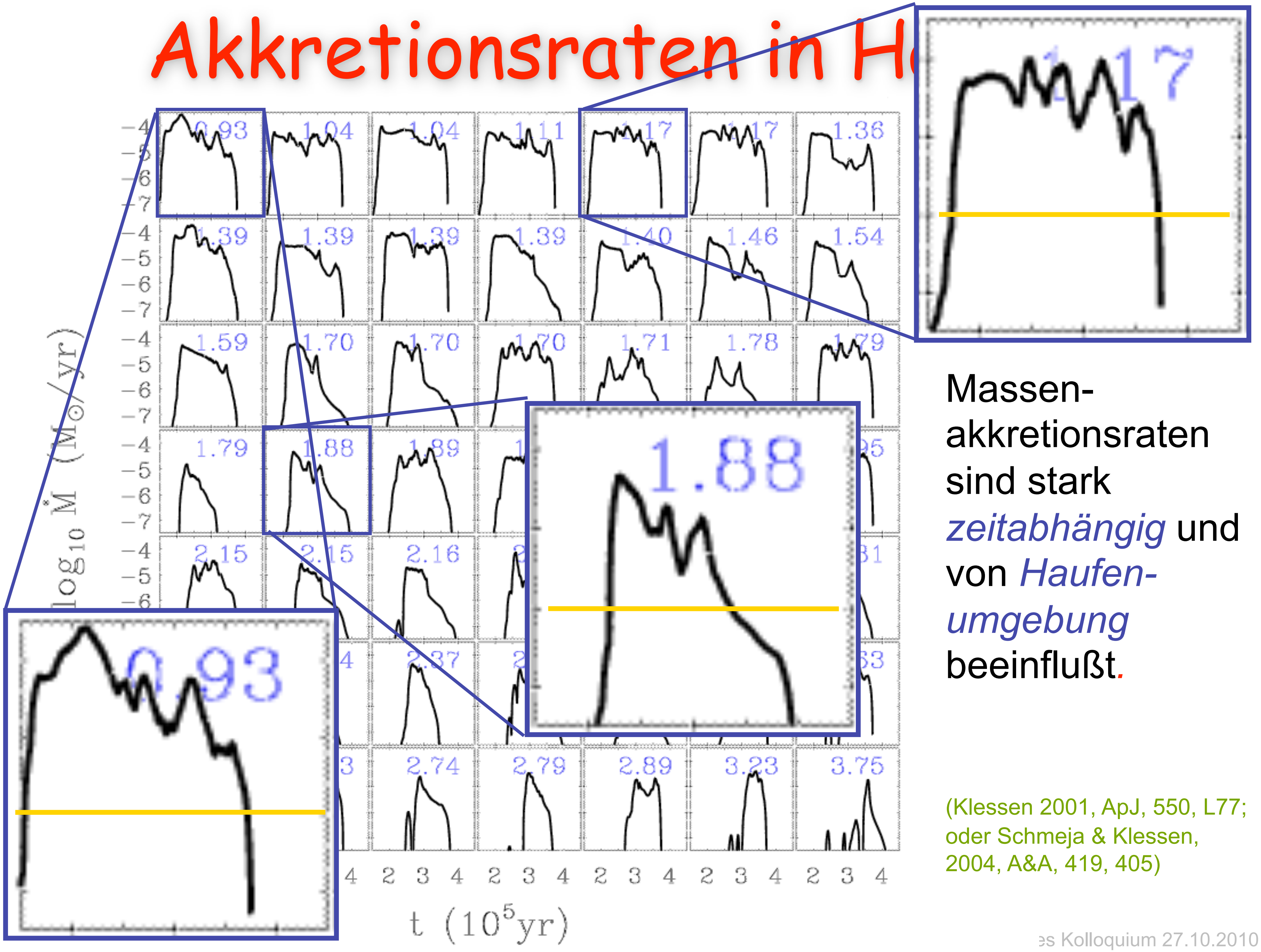
Bildung von Sternhaufen

Fast alle Sterne bilden sich in Haufen - dynamische Prozesse wichtig



Trajektorien junger Protosterne im sich bildenden Haufen.
(from Klessen & Burkert 2000, ApJS, 128, 287)

Akkretionsraten in H



Massen-akkretionsraten sind stark *zeitabhängig* und von *Haufen-umgebung* beeinflußt.

(Klessen 2001, ApJ, 550, L77; oder Schmeja & Klessen, 2004, A&A, 419, 405)

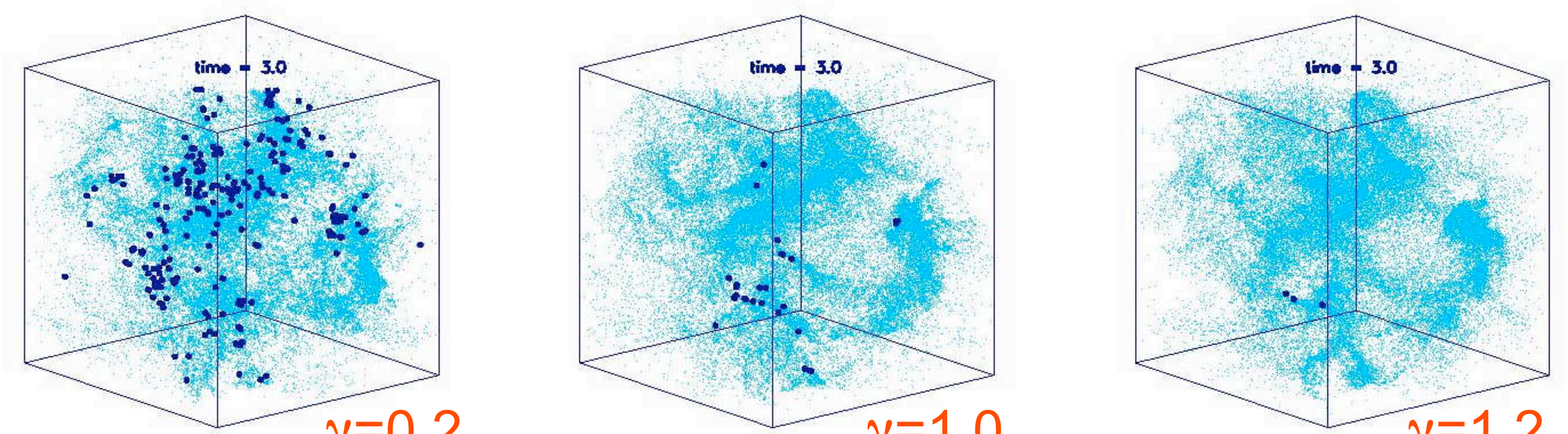
Abhängigkeit von Zustandsgl.

- ZG bestimmt Stärke der Fragmentation*!*

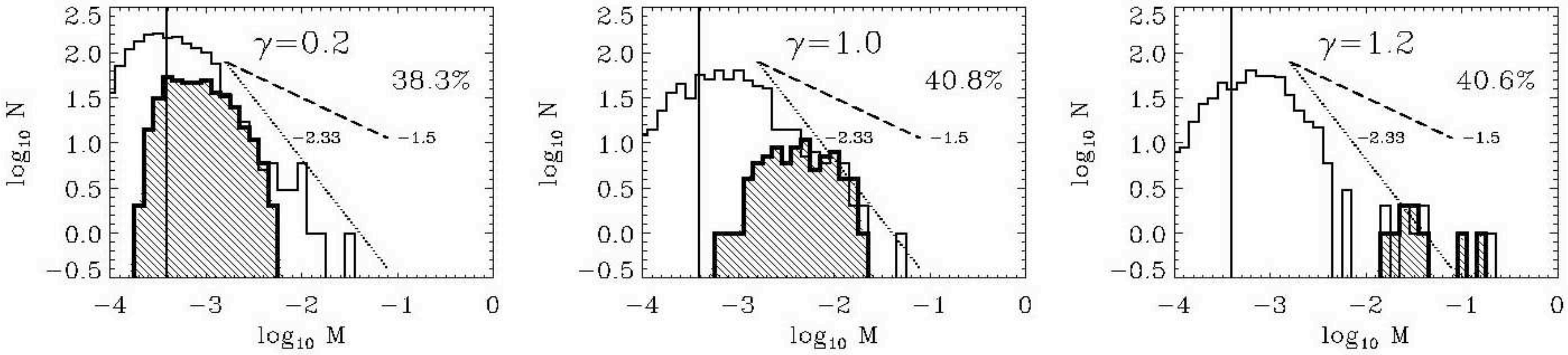
- polytrope ZG: $\mathbf{p} \propto \rho^{\gamma}$
- $\gamma<1$: dichter Haufen von massearmen Sternen
- $\gamma>1$: isolierte massereiche Sterne

(see Li, Klessen, & Mac Low 2003, ApJ, 592, 975; also Kawachi & Hanawa 1998, Larson 2003)

Abhängigkeit von ZG



γ=0.2 γ=1.0 γ=1.2



für $\gamma<1$ verstärkte Fragmentation → *Haufen massearmer Sterne*

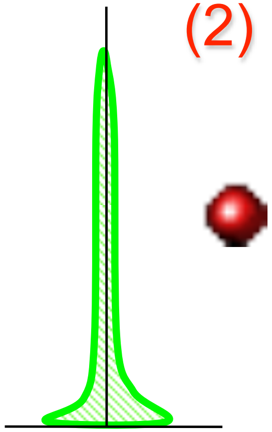
für $\gamma>1$ Fragmentation unterdrückt → Bildung *isolierter massereicher Sterne*

(from Li, Klessen, & Mac Low 2003, ApJ, 592, 975)

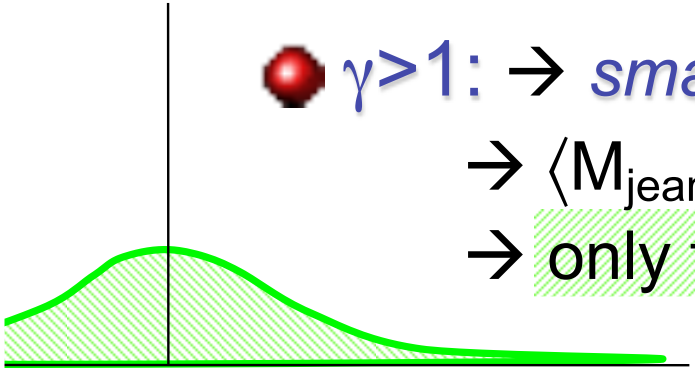
Wie funktioniert das?

(1) $\mathbf{p} \propto \rho^{\gamma}$ → $\rho \propto \mathbf{p}^{1/\gamma}$

(2) $\mathbf{M}_{\mathbf{jeans}} \propto \gamma^{3/2} \rho^{(3\gamma-4)/2}$



- $\gamma<1$: → *large* density excursion for given pressure
  → $\langle M_{jeans} \rangle$ becomes small
  → number of fluctuations with $M > M_{jeans}$ is large

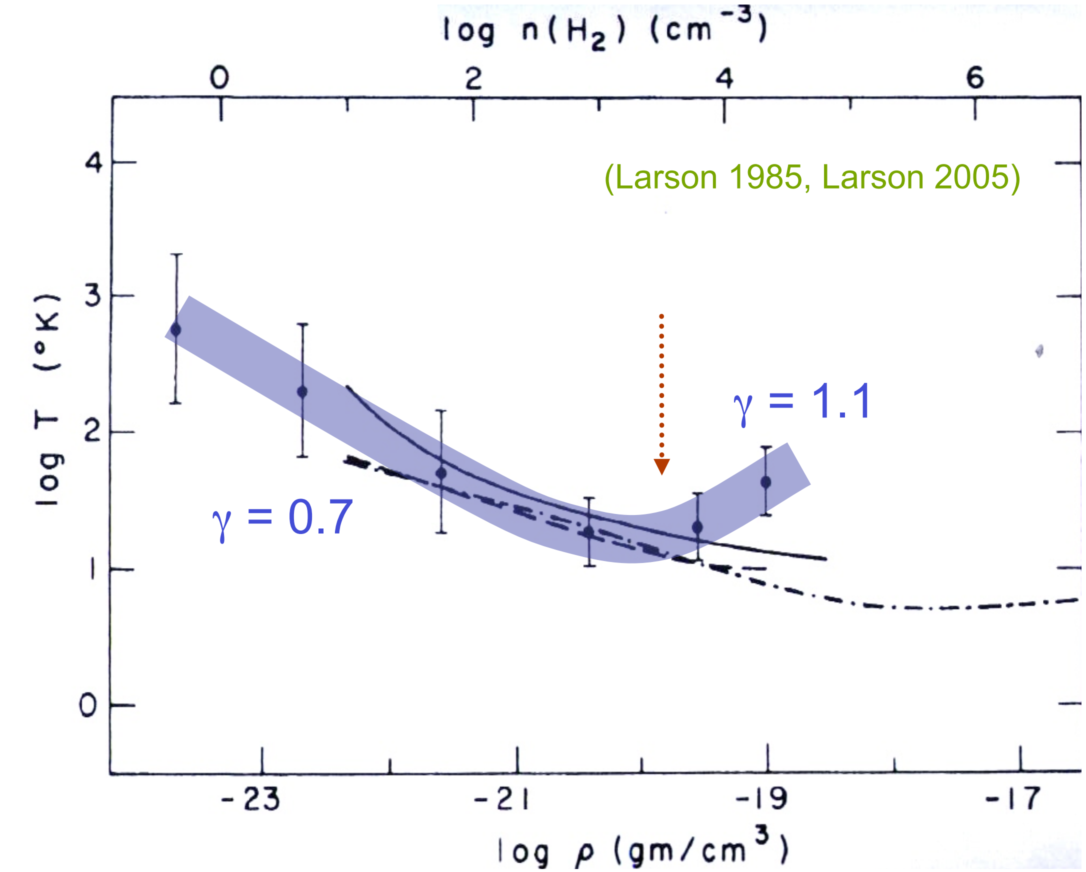


- $\gamma>1$: → *small* density excursion for given pressure
  → $\langle M_{jeans} \rangle$ is large
  → only few and massive clumps exceed M_{jeans}

Zustandsgleichung in Sonnennähe

unterhalb 10^{-18} $\mathrm{gcm^{-3}}$: ↑ρ ⟹ ↓T

oberhalb 10^{-18} $\mathrm{gcm^{-3}}$: ↑ρ ⟹ ↑T



(Larson 1985, Larson 2005)

$$P \propto \rho^{\gamma}$$

$$P \propto \rho T$$

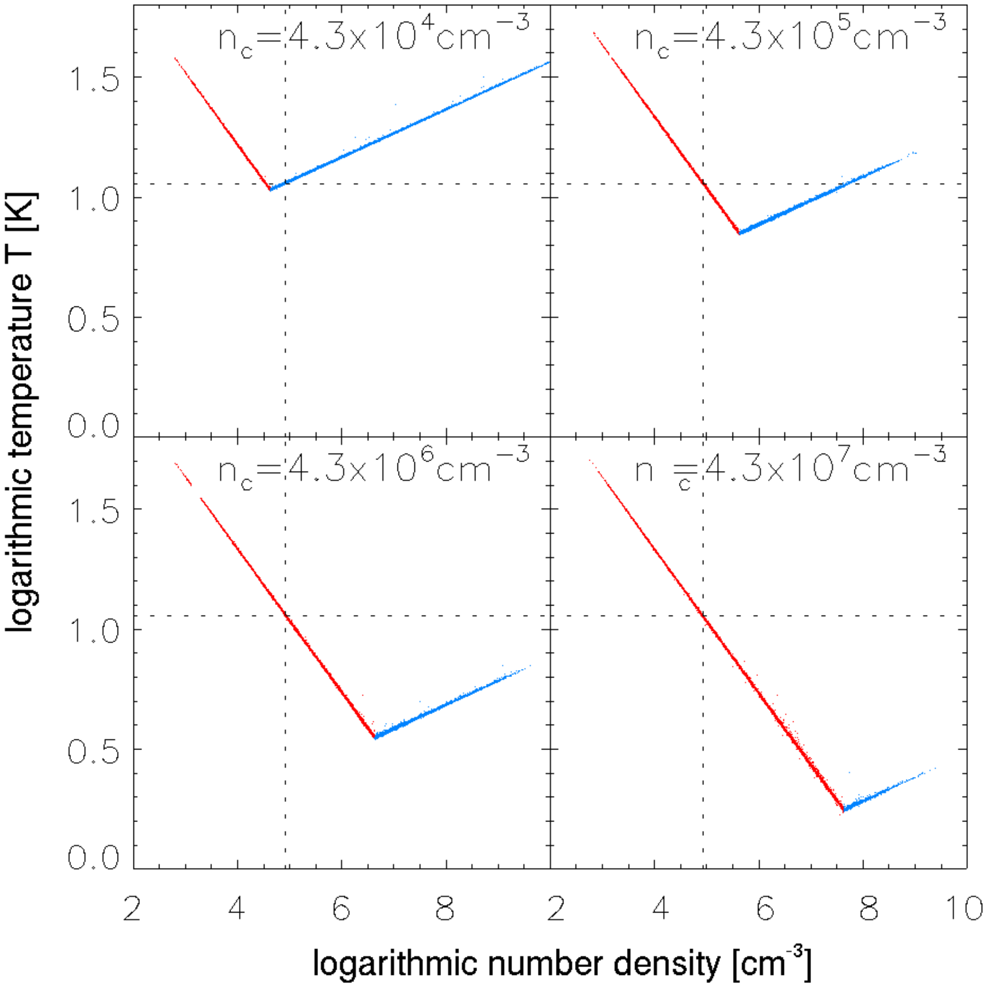
$$\rightarrow \gamma = 1 + d\ln T / d\ln \rho$$

Schrittweise polytrope ZG

$\gamma_1 = 0.7$

$\gamma_2 = 1.1$

$$\boxed{T \sim \rho^{\gamma-1}}$$



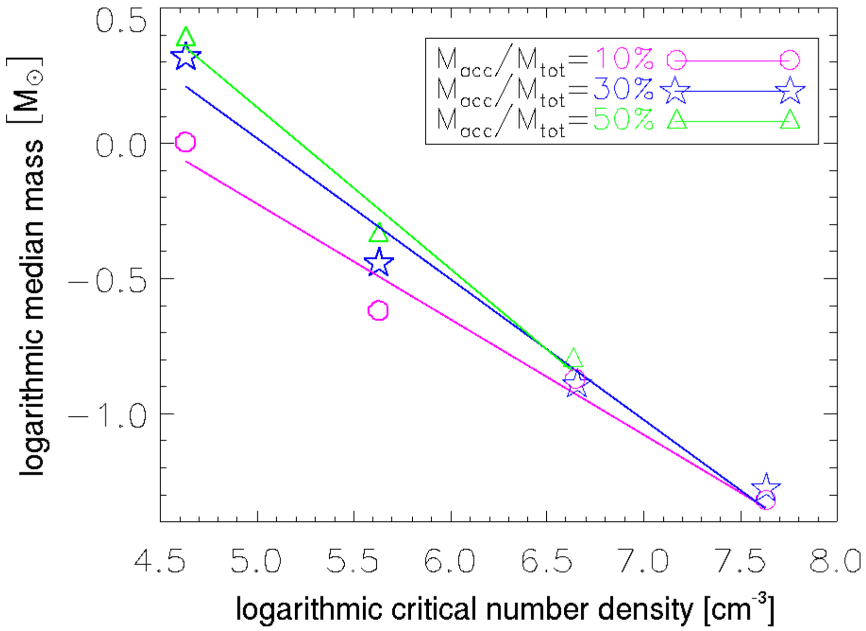
Simples numerisches Modell mit schrittweise polytroper Zustands-gleichung

(Jappsen et al. 2005)

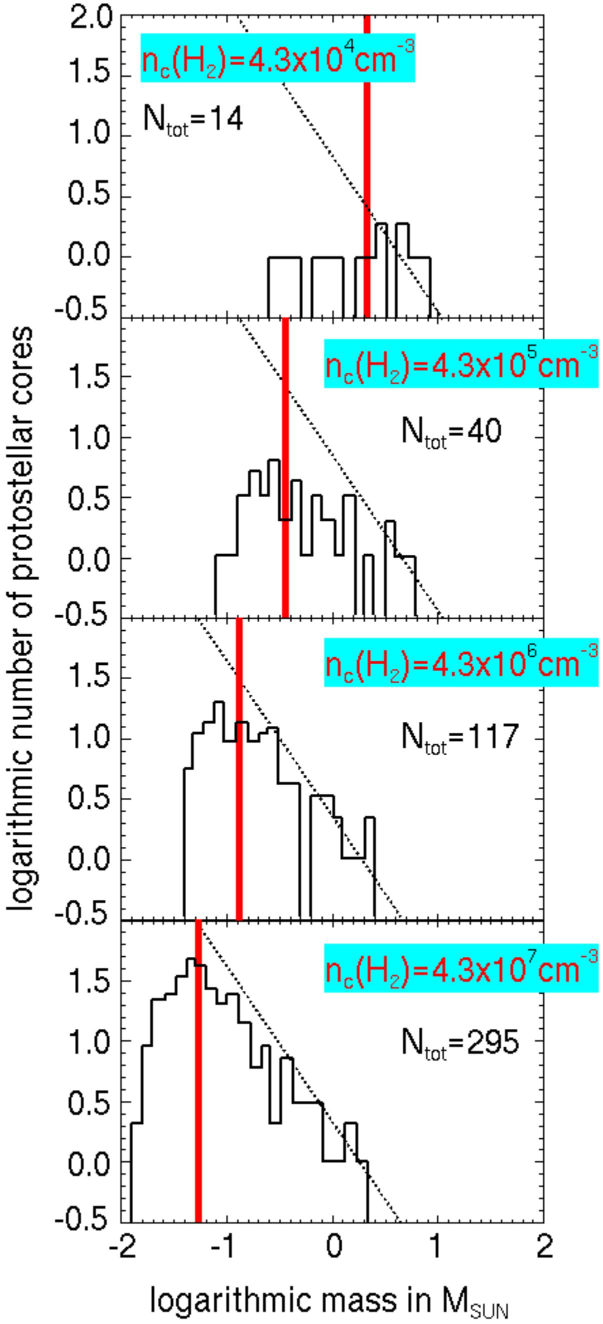
Stellare Massen-

kritische Dichte ↑

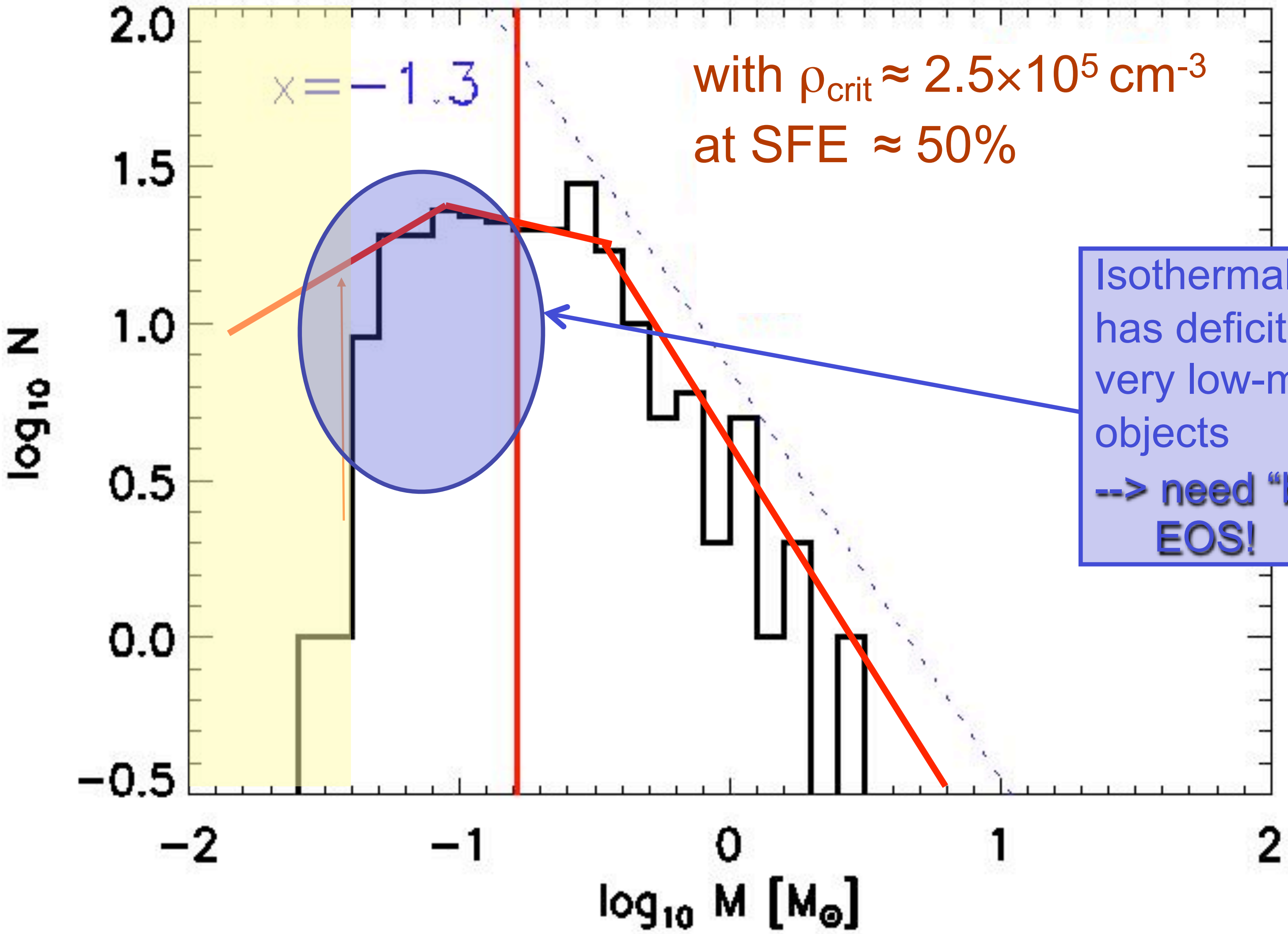
⇒ Median der Masse ↓



(Jappsen et al. 2005)



Vergleich mit beobachteter IMF



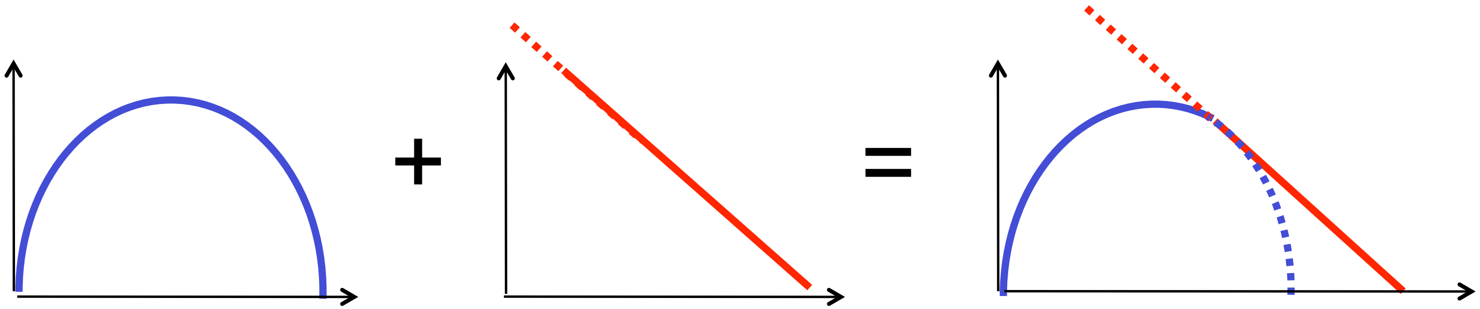
with $\rho_{crit} \approx 2.5\times10^5$ cm^{-3}
at SFE ≈ 50%

Isothermal EOS has deficite of very low-mass objects

--> need "better" EOS!

(Jappsen et al. 2005, A&A, 435, 611)

Plausibilitätsargument



- Überschallturbulenz ist skalenfreier Prozess

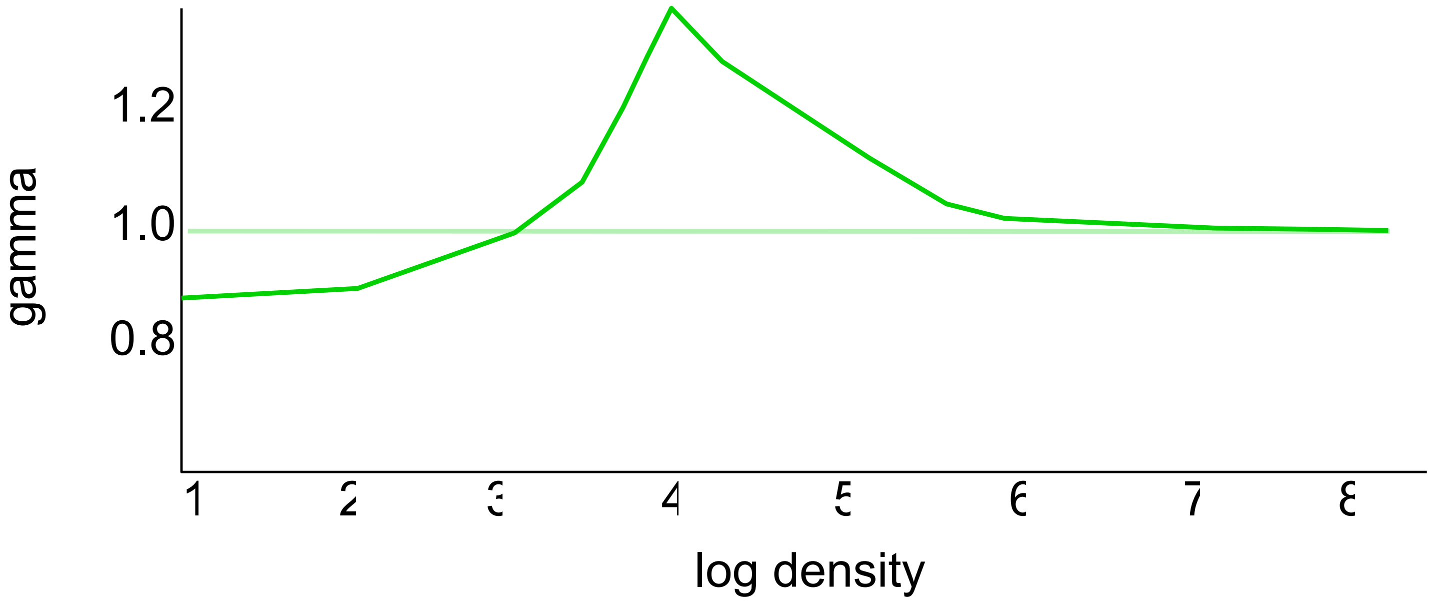
  → *Potenzgesetz*

- *Aber auch:* Turbulenz und Fragmentation sind hochgradig stochastische Prozesse → zentraler Grenzwertsatz
  → *Gauß Verteilung*

IMF in starburst galaxies

- Nuclear regions of starburst galaxies are extreme:
  - hot dust, large densities, strong radiation, etc.
- Thermodynamic properties of star-forming gas differ from Milky Way --> Different EOS!

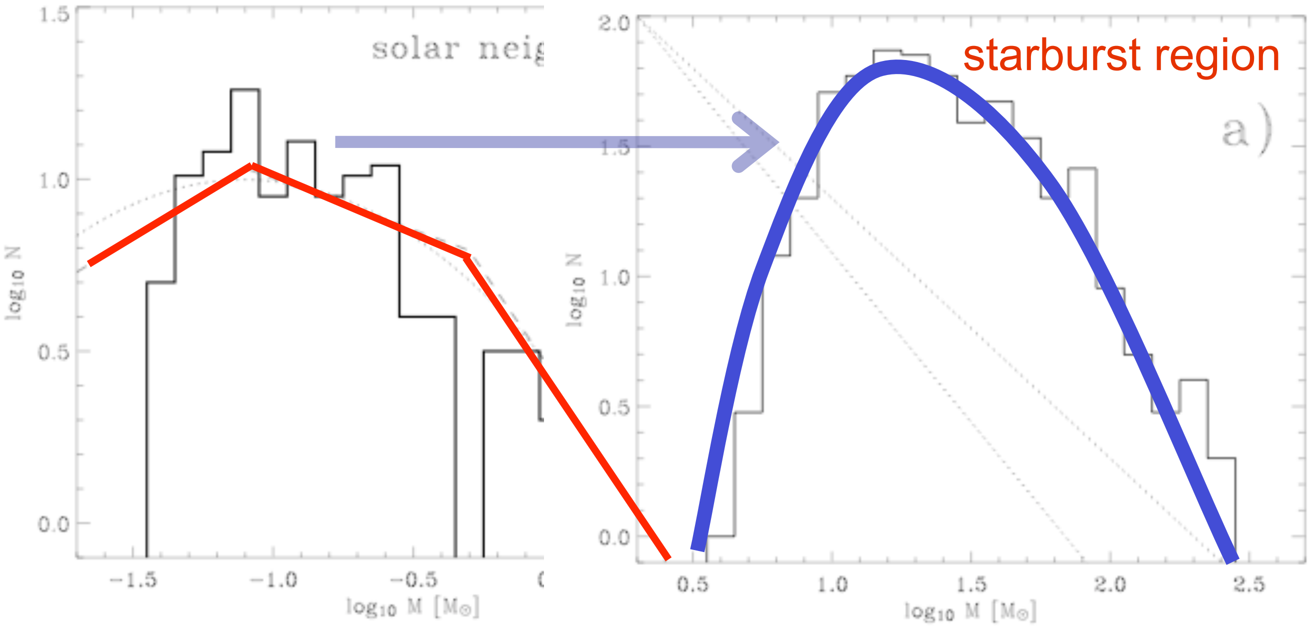
(see Spaans & Silk 2005)



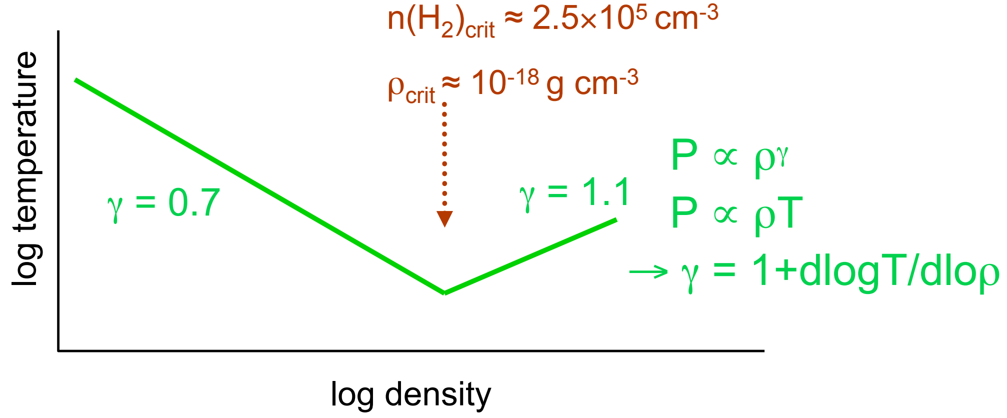
IMF in starburst galaxies

- Starburst EOS --> top-heavy IMF

(Klessen, Spaans, Jappsen, 2007)



fragmentation depends on EOS



$n(H_2)_{crit} \approx 2.5\times10^5\ \mathrm{cm}^{-3}$

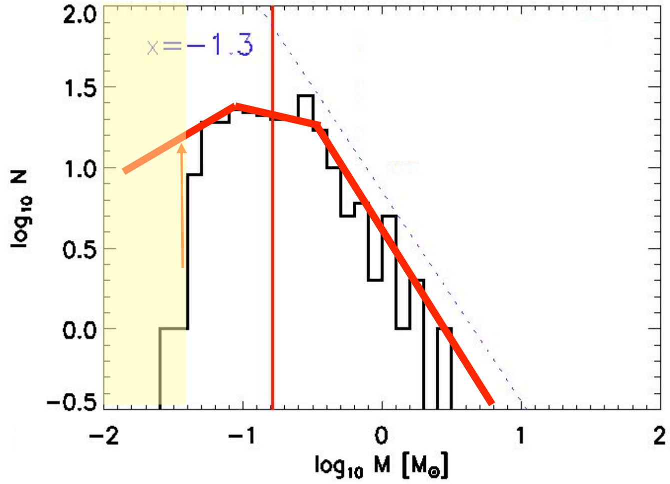
$\rho_{crit} \approx 10^{-18}\ \mathrm{g\ cm}^{-3}$

$P \propto \rho^\gamma$

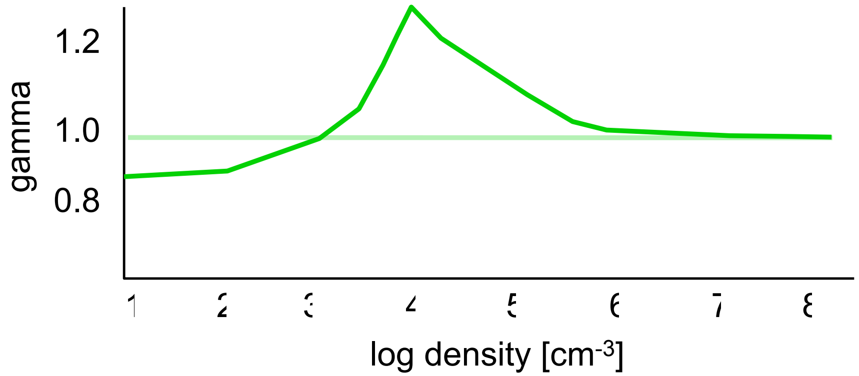
$P \propto \rho T$

$\rightarrow \gamma = 1 + d\log T / d\log\rho$

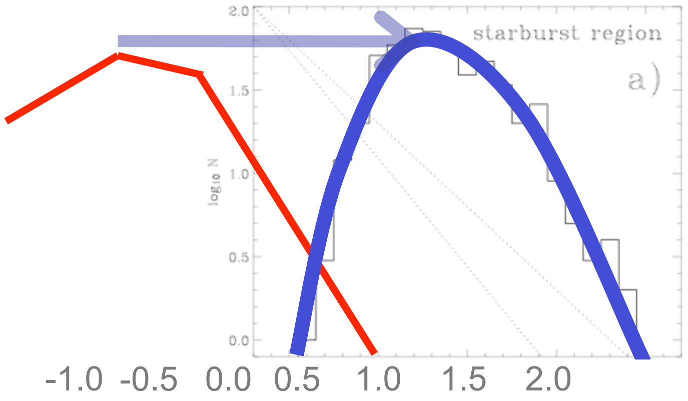
(Larson 2005)



(Jappsen et al. 2005)

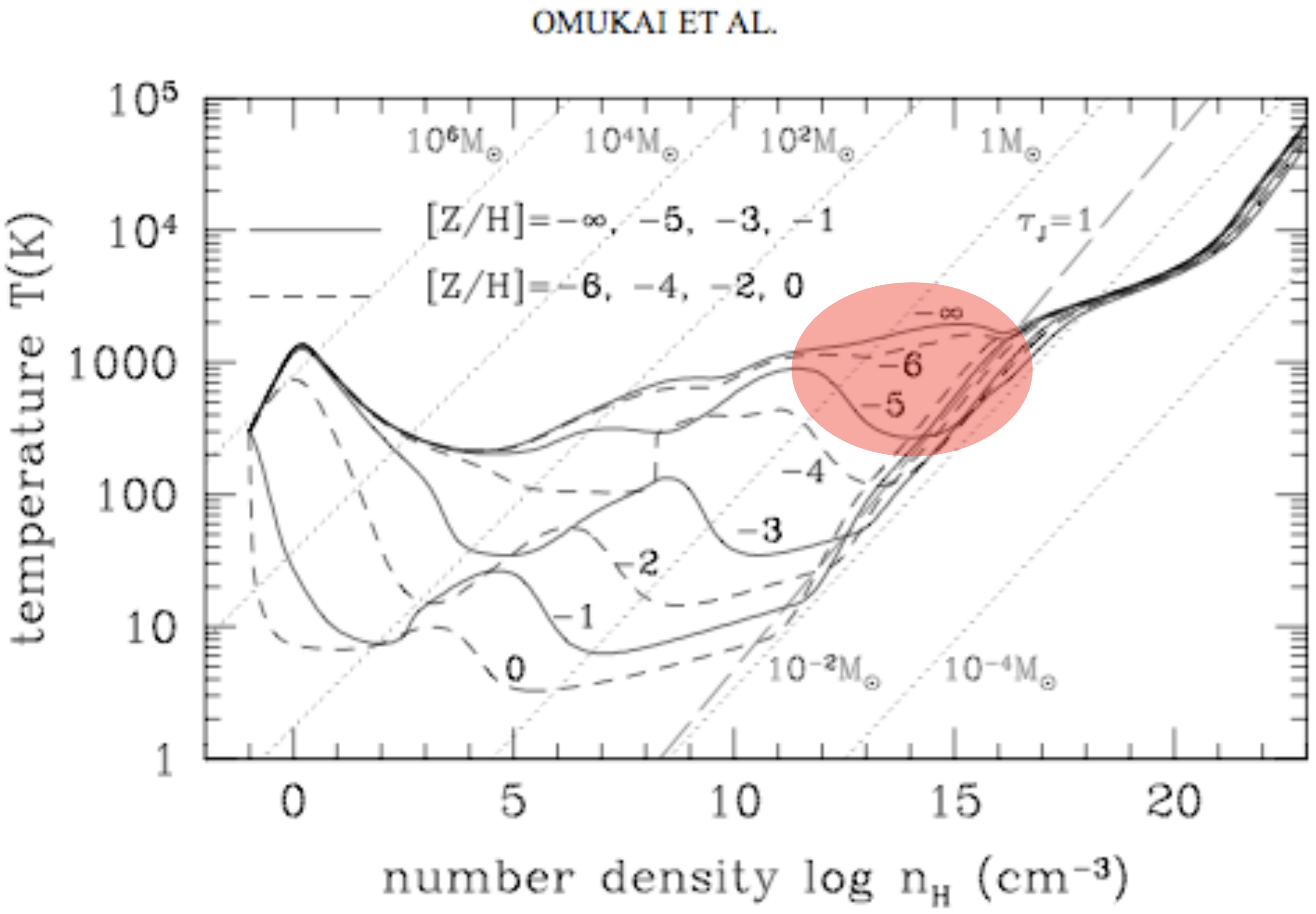


(Spaans & Silk 2005)



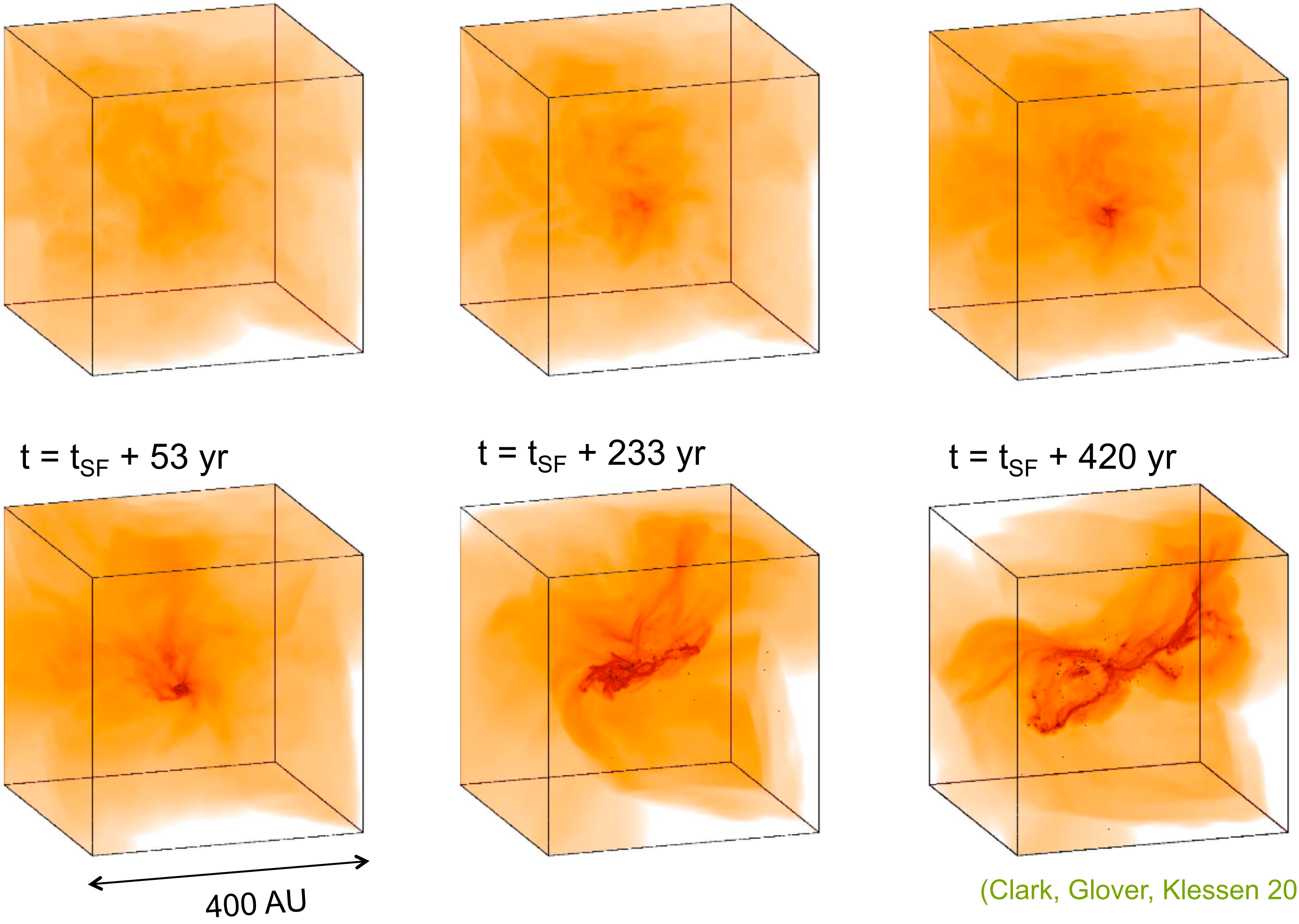
(Klessen et al. 2007)

Übergang: Pop III --> Pop II.5



(Omukai et al. 2005)

Staubinduzierte Frag. bei $Z=10^{-5}$



$t = t_{SF} - 67$ yr

$t = t_{SF} - 20$ yr

$t = t_{SF}$

$t = t_{SF} + 53$ yr

$t = t_{SF} + 233$ yr

$t = t_{SF} + 420$ yr

(Clark, Glover, Klessen 2007)

Staubinduzierte Frag. bei $Z=10^{-5}$

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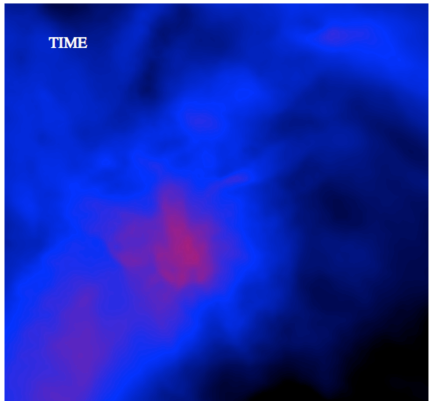
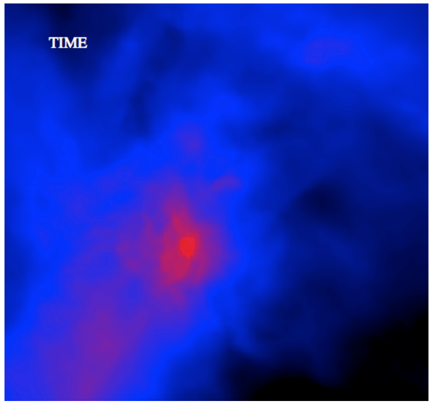
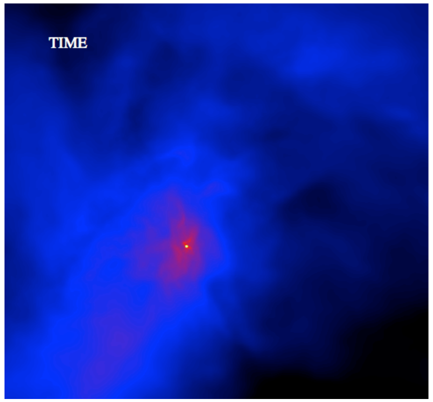


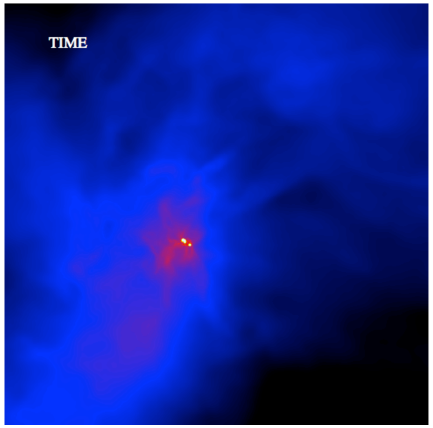
$t = t_{SF} - 20$ yr



$t = t_{SF}$



$t = t_{SF} + 53$ yr



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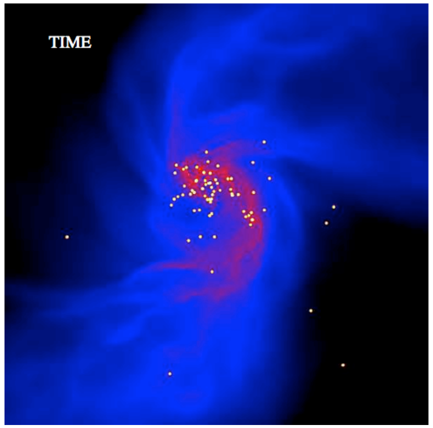

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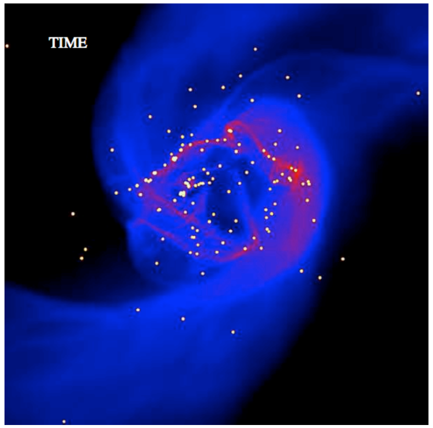

(Clark et al. 2007

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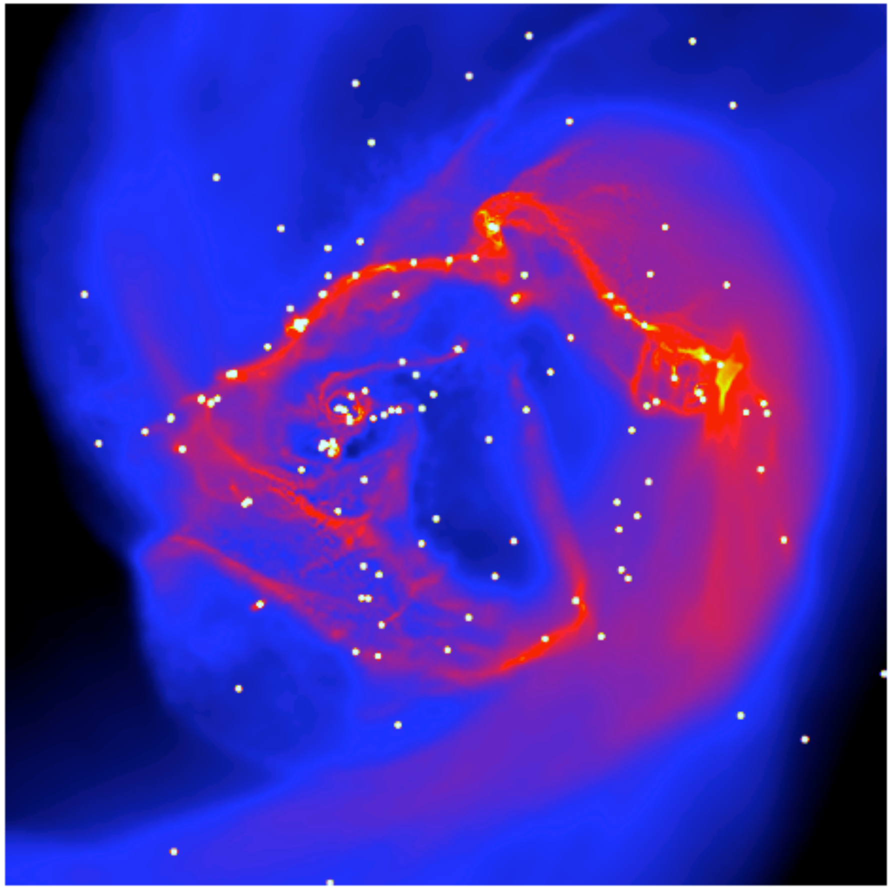


extrem dichter Haufen massearmer Sterne:

- Maximum der IMF unterhalb von 1 M_{sun}
- Haufen SEHR dicht $n_{stars} = 2.5 \times 10^9\ pc^{-3}$
- Fragmentation bei Dichten von $n_{gas} = 10^{12} - 10^{13}\ cm^{-3}$

(Clark et al. 2007)

Aufbau des Sternhaufens

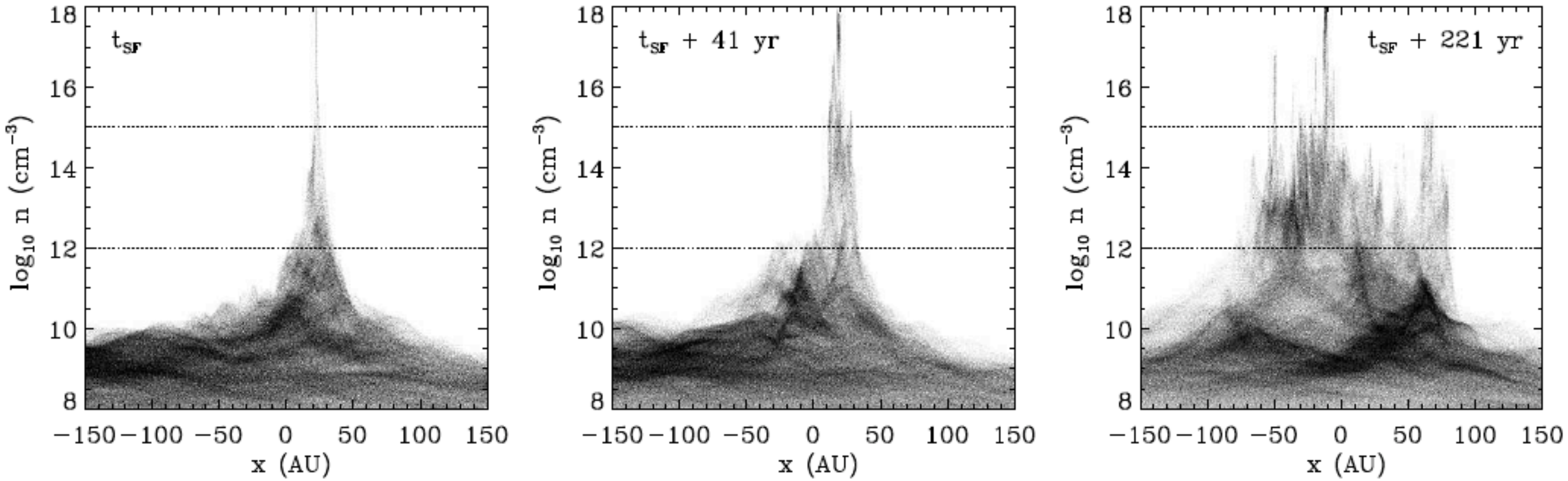


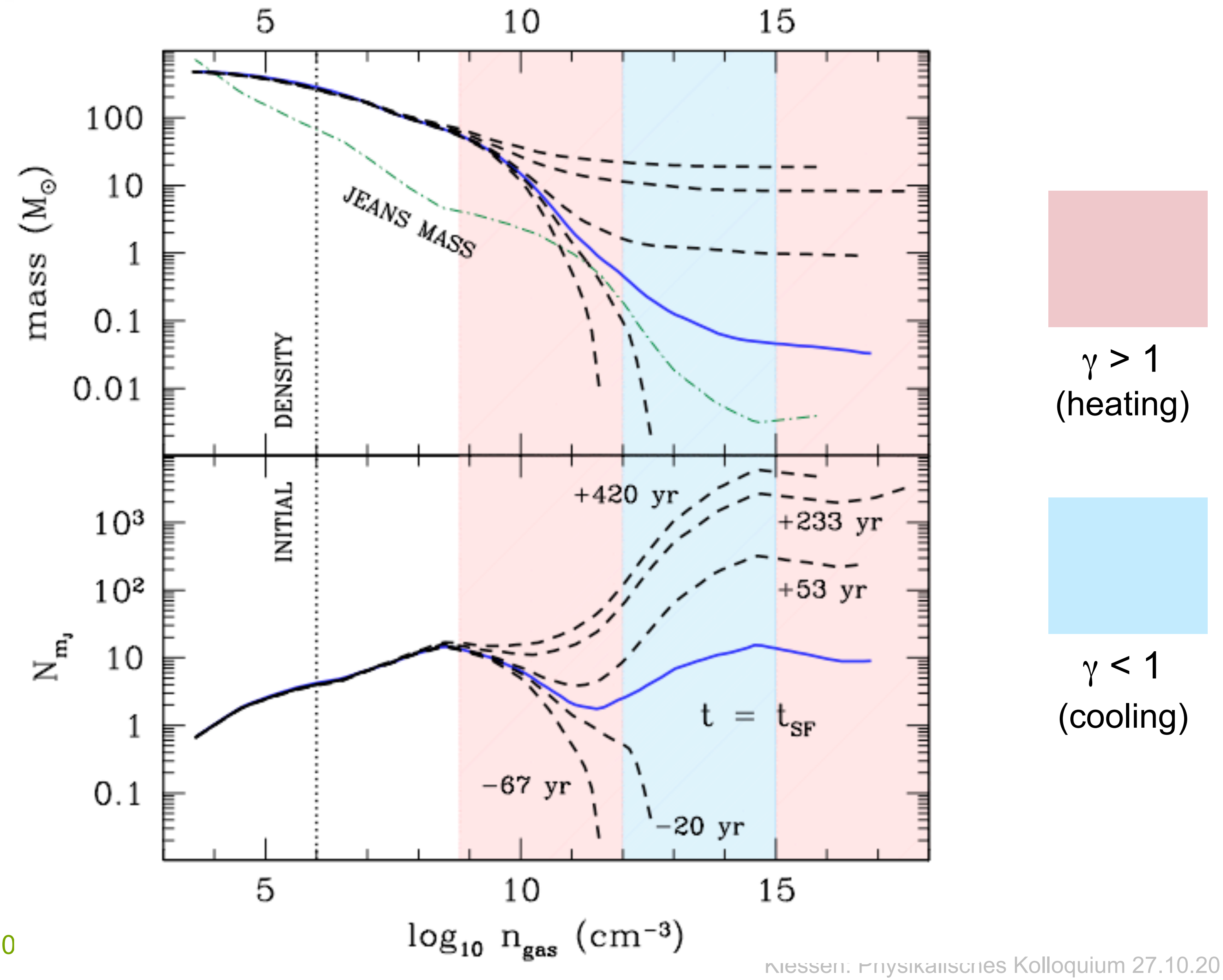
FIG. 3.— We illustrate the onset of the fragmentation process in the high resolution $Z = 10^{-5}\,Z_\odot$ simulation. The graphs show the densities of the particles, plotted as a function of their x-position. Note that for each plot, the particle data has been centered on the region of interest. We show here results at three different output times, ranging from the time that the first star forms (t_{sf}) to 221 years afterwards. The densities lying between the two horizontal dashed lines denote the range over which dust cooling lowers the gas temperature.

→

Zeit

(Clark et al. 2007)

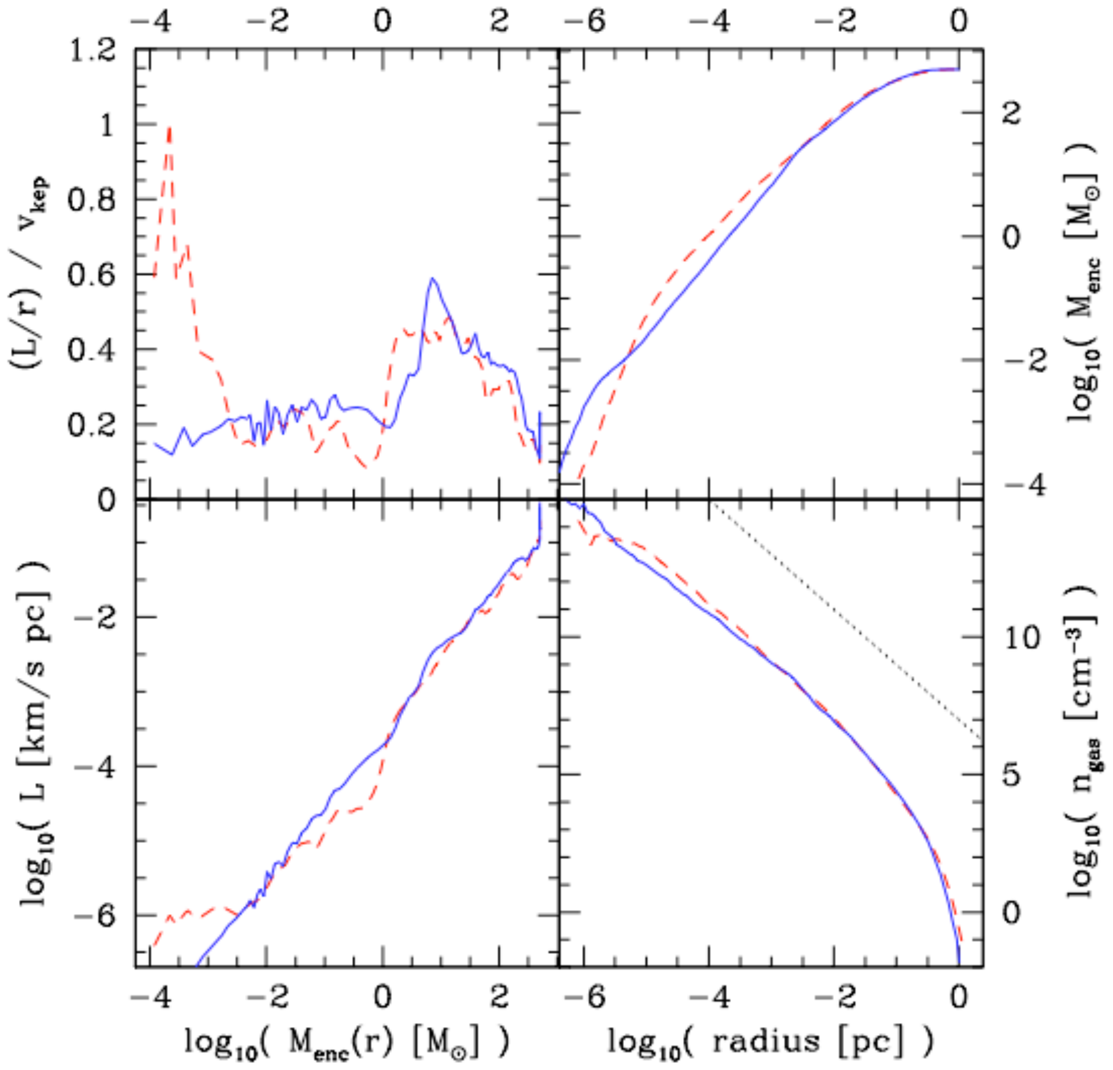
Aufbau des Sternhaufens



(Clark et al. 20

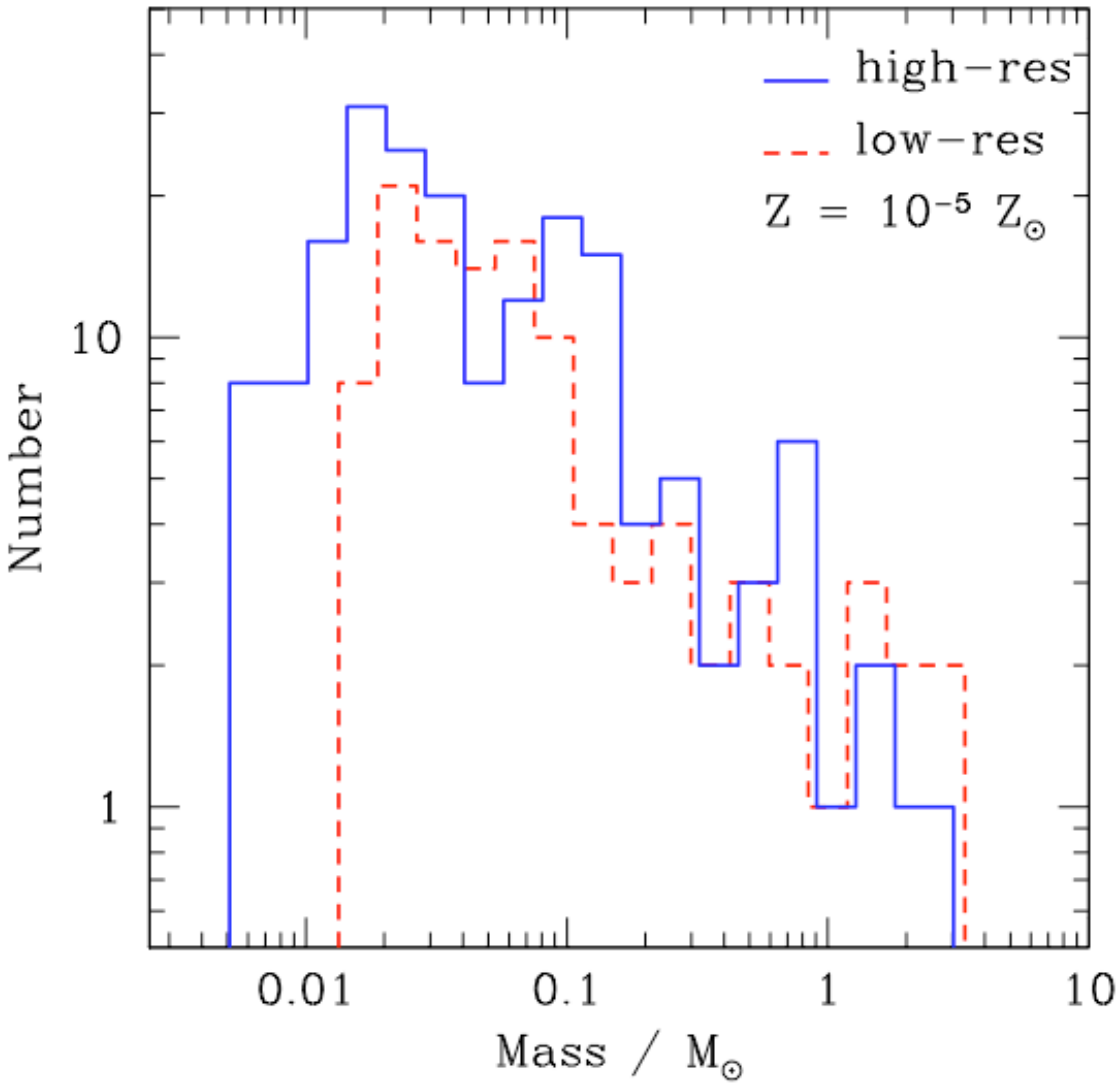
Eigenschaften des Gases

Gaseigenschaften zum Zeitpunkt der Bildung des ersten Sternes



(Clark et al. 2007)

Staubinduzierte Frag. bei $Z=10^{-5}$



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Vergleiche unterschiedliche Z

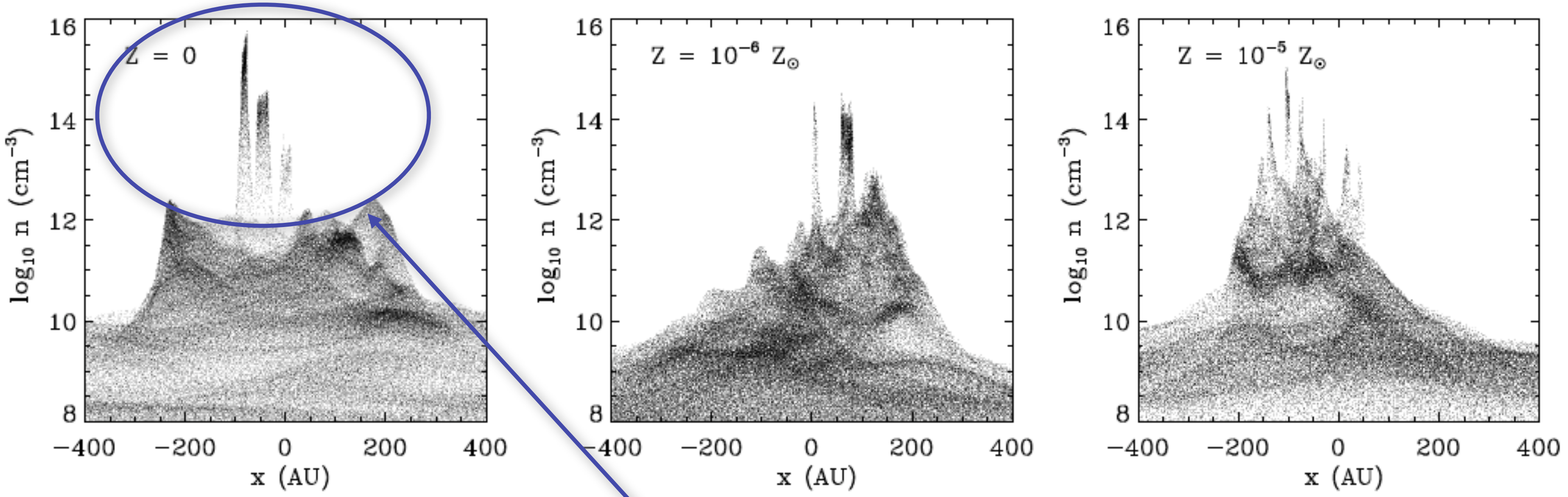

FIG. 6.— Particle densities as a function of position in the low-resolution simulations, for the primordial (left), $Z = 10^{-6}\,Z_\odot$ (middle) and $Z = 10^{-5}\,Z_\odot$ simulations (right). The particles are plotted once the protostars in each simulation have accreted 19 $M_\odot$ of gas.

Selbst metallfreies Gas fragmentiert
(wenn auch wesentlich schächer) !

Unterschied zwischen
Pop III.1 & Pop III.2 ?

(Clark et al. 2007)

Vergleiche unterschiedliche Z

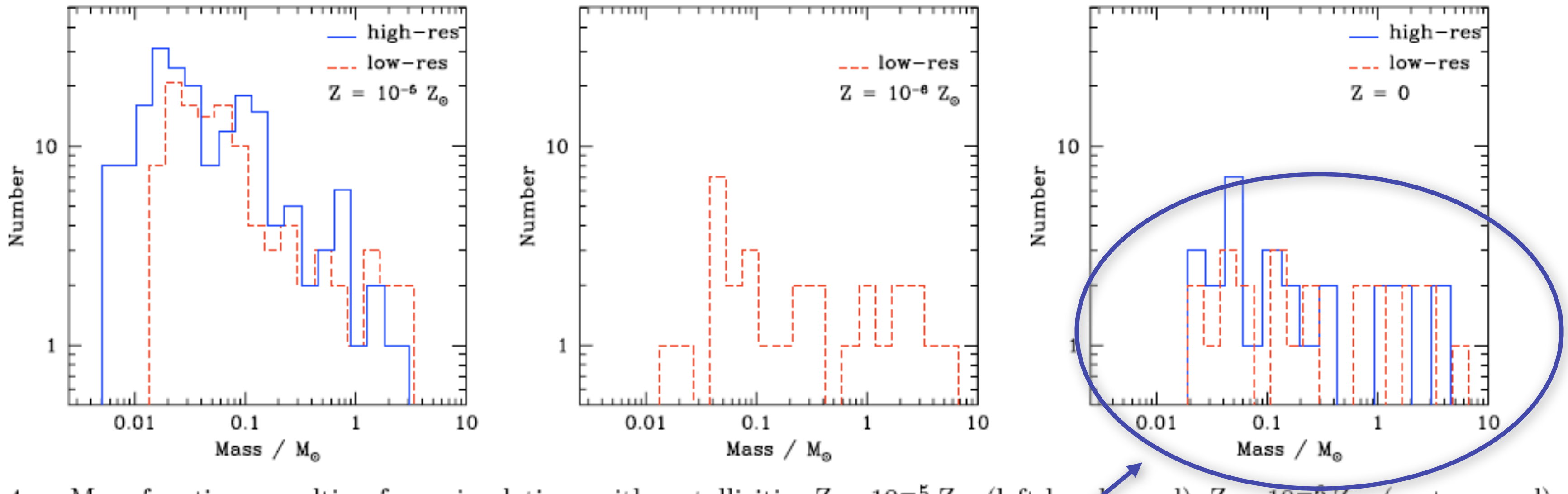

FIG. 4.— Mass functions resulting from simulations with metallicities $Z = 10^{-5}\,Z_\odot$ (left-hand panel), $Z = 10^{-6}\,Z_\odot$ (center panel), and $Z = 0$ (right-hand panel). The plots refer to the point in each simulation at which 19 $M_\odot$ of material has been accreted (which occurs at a slightly different time in each simulation). The mass resolutions are 0.002 $M_\odot$ and 0.025 $M_\odot$ for the high and low resolution simulations, respectively. Note the similarity between the results of the low-resolution and high-resolution simulations. The onset of dust-cooling in the $Z = 10^{-5}\,Z_\odot$ cloud results in a stellar cluster which has a mass function similar to that for present day stars, in that the majority of the mass resides in the lower-mass objects. This contrasts with the $Z = 10^{-6}\,Z_\odot$ and primordial clouds, in which the bulk of the cluster mass is in high-mass stars.

Selbst metallfreies Gas fragmentiert
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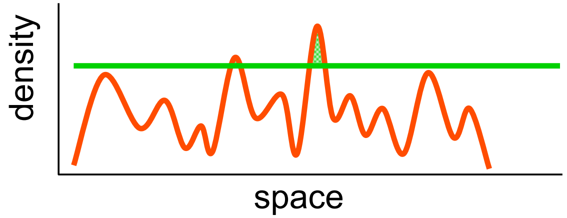
(Clark et al. 2007)

Zusammenfassung

Summary I



- interstellar gas is highly *inhomogeneous*
  - *thermal instability*
  - *gravitational instability*
  - *turbulent compression* (in shocks $\delta\rho/\rho \propto \mathcal{M}^2$; in atomic gas: $\mathcal{M} \approx 1...3$)
- cold *molecular clouds* can form rapidly in high-density regions at *stagnation points* of *convergent large-scale flows*
  - chemical *phase transition*: atomic → molecular
  - process is *modulated* by large-scale *dynamics* in the galaxy
- inside *cold clouds:* turbulence is highly supersonic ($\mathcal{M} \approx 1...20$)
  → *turbulence* creates density contrast, *gravity* selects for collapse

⟶ ***GRAVOTUBULENT FRAGMENTATION***

- *turbulent cascade:* local compression *within* a cloud provokes collapse → formation of individual *stars* and *star clusters*
- *star cluster:* gravity dominates in large region (--> competitive accretion)

(e.g. Mac Low & Klessen, 2004, Ballesteros-Paredes et al. 2006, McKee & Ostriker 2007)

Summary II

- *thermodynamic response* (EOS) determines fragmentation behavior
  - characteristic stellar mass from fundamental atomic and molecular parameters
    --> explanation for quasi-universal IMF?

- *stellar feedback* is important
  - accretion heating may reduce degree of fragmentation
  - ionizing radiation will set efficiency of star formation
- *CAVEATS:*
  - star formation is *multi-scale, multi-physics* problem --> VERY difficult to model
  - in simulations: very small turbulent inertial range (Re < 1000)
  - can we use EOS to describe thermodynamics of gas, or do we need time-dependent chemical network and radiative transport?
  - stellar feedback requires (at least approximative) radiative transport, most numerical calculations so far have neglected that aspect

Sternentstehung

- *Sterne entstehen in überschall-turbulenten Molekülwolken.*
- *Interstellare Turbulenz bestimmt Dichte und Geschwindigkeitsstruktur*
- *Protosterne bilden sich durch gravoturbulent Fragmentation*
- *Fast alle Sterne bilden sich in Haufen.* → *Wechselwirkung in der Hauptakkretionsphase*
- *Sternentstehung ist hochgradig stochastischer Prozess*
- *Zeitskala und räumliche Verteilung der Sterne bestimmt durch ISM Turbulenz*
- *Thermodynamik des Gases bestimmt Fragmentationseigenschaften (fundamentale atomare und molekulare Parameter)*
- *Universelles stellares Massenspektrum (Abweichungen im frühen Universum)*

Ausblick

- Komplettes Bild: *Verbinde die (lokale) Sternentstehung mit der globalen dynamischen Entwicklung der Galaxie!*
  - Wie *beeinflussen* globale Prozesse die SE auf kleinen Skalen? Deren Summe wiederum ergibt die *globale* SE-Geschichte der Galaxie
- Und umgekehrt: *Wie beeinflusst SE auf kleinen Skalen die globalen Eigenschaften der Milchstraße?*
  - Was treibt Turbulenz im interstellaren Medium?
  - Welche *Mischungseigenschaften* hat ISM Turbulenz → wichtig für das Verständnis der chemischen Entwicklung der Milchstraße
  - Galactic Fountains, Materiekreislauf, Schmidt Gesetz, Relation Galaxienmasse - zentrales Schwarzes
- *Wann, wo und wie bilden sich die erste und zweite Sterngeneration (Pop III --> Pop II.5)*
  - Welche Masse haben Pop III? Welche Rolle spielen Umwelteinflüsse?
  - Kosmische Reionisation, Bildung der ersten Galaxien

Danke!